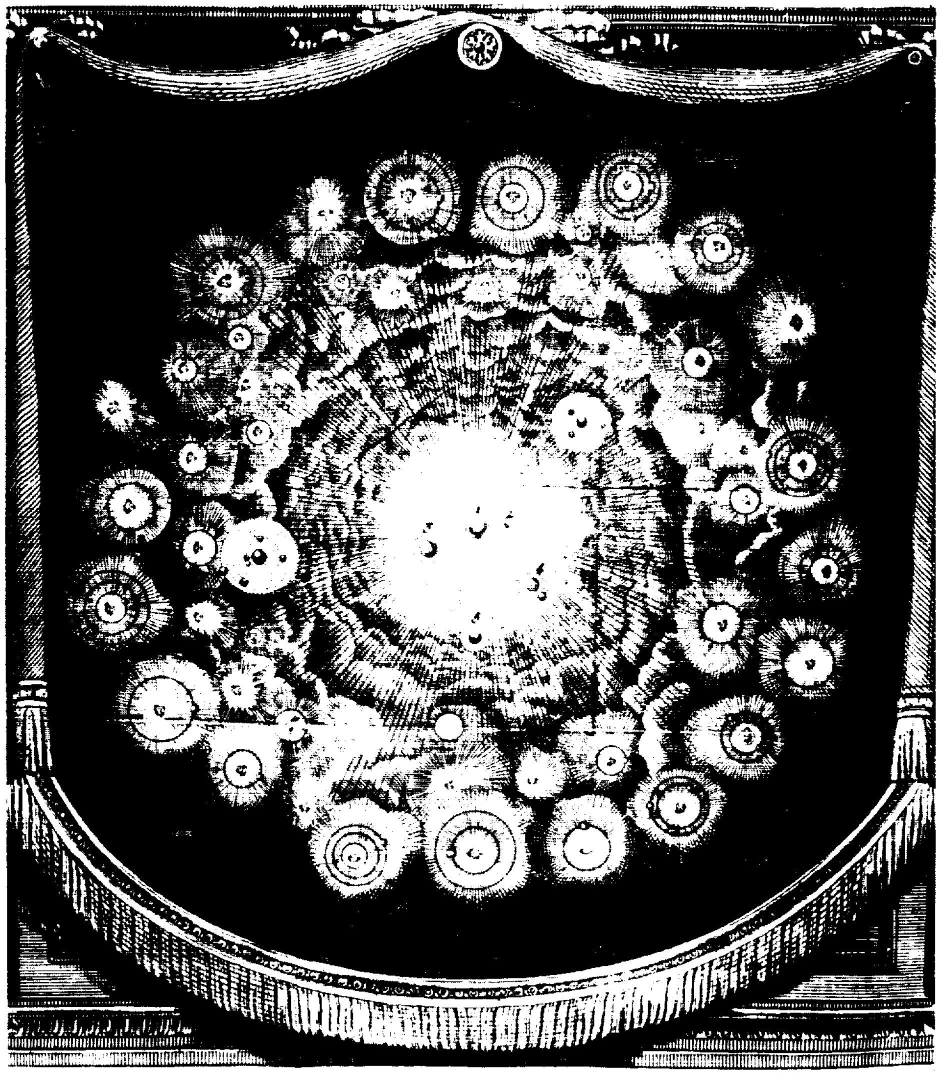
NASA SP-497

O STARS AND WOLF-RAYET STARS

NOGRAPH SERIES ON NONTHERMAL PHENOMENA
IN STELLAR ATMOSPHERES

O STARS AND WOLF-RAYET STARS



1 Le Soleil 2 Mercure 3 Venus 4 La Terre 5 Mars 6 Jupiter 7 Saturne

... Assurons-nous bien du fait, avant que de nous inquiéter de la cause. Il est vrai que cette méthode est bien lente pour la plupart des gens, qui courent naturellement à la cause, et passent par dessus la vérité du fait; mais enfin nous éviterons le ridicule d'avoir trouvé la cause de ce qui n'est point.

...

...

... De grands physiciens ont fort bien trouvé pourquoi les lieux souterrains sont chauds en hiver, et froids en été; de plus grands physiciens ont trouvé depuis peu que cela n'était pas.

–Fontenelle, Histoire des Oracles
Chapitre IV, pp. 20 et 23

NASA SP-497

MONOGRAPH SERIES ON NONTHERMAL PHENOMENA
IN STELLAR ATMOSPHERES

O STARS AND WOLF-RAYET STARS

Peter S. Conti
Anne B. Underhill

Stuart Jordan, Solar Physics
Editor/Organizer and American Coordinator

Richard Thomas, Stellar/Solar
Editor/Organizer and European Coordinator

Leo Goldberg
Senior Adviser to NASA

Jean-Claude Pecker
Senior Adviser to CNRS

Centre National de la
Recherche Scientifique
Paris, France

1988

National Aeronautics and
Space Administration
Scientific and Technical
Information Division
Washington, D.C.

LIBRARY OF CONGRESS
Library of Congress Cataloging-in-Publication Data

O-type stars and Wolf-Rayet stars / [edited by] Peter S. Conti, Anne B. Underhill.
p. cm. -- (Monograph series on nonthermal phenomena in stellar atmospheres) (NASA SP-497)
1. O stars. 2. Wolf-Rayet stars. I. Conti, P. S. (Peter S.) II. Underhill, Anne Barbara, 1920- . III. Series. IV. Series: NASA SP ; 497.
QB842.012024 1988
523.8'7--dc19 88-15160
CIP

For sale by the Superintendent of Documents,
U.S. Government Printing Office, Washington, DC 20402

DEDICATION

in grateful appreciation
we dedicate this series and these volumes

to *Cecilia Payne-Gaposchkin*, who, with Sergei, set the spirit of empirical-theoretical atmospheric modeling by observing:

> "All true variable stars have variable atmospheres, but a variable atmosphere is probably the property of all stars, whether obviously variable in brightness or not [as witness the solar envelope]";

and who, by her intimate knowledge of particular stars, pioneered in the recognition of the fundamental importance of "individuality of stellar atmospheric characteristics."

to *Daniel Chalonge*, who sought, by ingenious meticulous observations, to make quantitative the features of qualitative classical taxonomy, thereby laying the foundations of showing the inadequacy of its two-dimensional, single-region atmospheric, character;

and who always opposed the spirit of a distinguished theoretical colleague's remark:

> "Don't show me those new observations of yours; they inhibit the range of my speculations."

LEO GOLDBERG

in Memoriam

Leo died in November 1987, ten days before we received the M-Star volume from the printer. Leo was a working, nonpolitical, senior advisor of this Monograph Series, as well as a long-time trusted colleague and friend. He was characteristically passionate toward the series objective—to delineate the phenomenon before modeling it—and dispassionate in his counsel on how to resolve those problems clouding the objective. He was characteristically visionary and innovative toward the series methodology—to bring together high-resolution solar/stellar, terrestrial, and spatial observations in order to give such delineation a panspectral breadth.

Each volume of this series contains his contribution to its development, in addition to his knowledge of, and curiosity about, the particular stars and their astrophysics. He labored with us, with the editors, and with the authors to shape this volume in the light of both its predecessors in the series and the continuous change in perspective coming from combined ground/space and stellar/solar observations. Such perspective was Leo's strength in his scientific career.

Leo had been acclaimed "Mr. Sun" among his astronomical colleagues for developing high-resolution ground-based solar facilities at Michigan, and then complementing their output with that from the solar satellite studies on which he focused Harvard's solar program on his return there as Director of the Observatory. This title was hardly pejorative to his mentor and predecessor in such identification—Donald Menzel—who, with his students, gave a new dimension to stellar astrophysics by exhibiting the common nonEquilibrium physics of the planetary nebulae and the outer solar atmosphere—hence inferentially to all the low-density atmospheres which embrace most of what we study as the frontier in nonEquilibrium nonthermal atmospheric phenomena today.

The Harvard of that era was more than just Shapley's carefully assembled interacting environment of Payne-Gaposchkin, Menzel, Whipple, and Bok; it was also strongly the first-generation students of each who complemented them in training and inspiring those of us who came a few years later. Leo was one of the trio "Aller, Baker, and Goldberg" who, with Menzel, formulated that first-approximation approach to gaseous nebulae adopted quasi-universally in studying a variety of low-density atmospheres. Leo went on to the Sun and cool stars, Aller to hot stars, and Baker to ingenious telescope designs. Another pair of students, Evans (under Bok) and Roberts (Menzel), formed the chief competitive collaboration with Leo's solar studies. With Menzel, the pair originally replaced Menzel's eclipse basis by Lyot's coronagraph, and then returned to Menzel's eclipse approach to launch the HAO-Sac Peak era of exploiting the low-density nonEquilibrium regions of stellar atmospheres via solar guidance. With Leo's accession to the AURA Directorship, those two branches of Harvard-inspired solar/stellar astronomy were at least partially reunited under one of their founders. Leo's allegiance to his Harvard origins and training were profound and reciprocal, even after he moved on.

PAGE VI INTENTIONALLY BLANK

PRECEDING PAGE BLANK NOT FILMED

As Director of KPNO, at a time when both SPO and CTIO were under his oversight, Leo provided incisive leadership by his visionary presence and his productive collaboration with the talented staff he recruited and supported. Goldberg believed vehemently that scientific leadership demands support of the most creative scientists who are responsible for the scientific excellence of the institution. He was adamantly opposed to any compromise in the quality of the scientific endeavor with which he was associated. He believed in institutional loyalty and had little patience with institutions managed by committees meeting often in secret and rife with conflicts of interest. One of his favorite quotations can be paraphrased: "It takes talent to recognize genius; mediocrity knows only itself."

The same joint focus on space plus ground observations which made him Mr. Sun to the astronomers made him "Mr. Space Astronomy" to NASA. He had long debated between a move to NASA/Goddard Space Flight Center, or to Kitt Peak National Observatory/Tucson, when he decided on the latter. It is ironic that the collaboration with Bd, begun with the AURA move to Tucson, returned him to NASA/space via their joint efforts, after he stepped down as Director, culminating in AURA taking on the ST. At the epoch when the game was "follow-the-leader" in planning large ground-based telescopes, Leo's focus lay on high-resolution by space-based interferometers. Again, high-resolution, visual and far ultraviolet, Sun and stars.

Leo brought the same sense of excellence and visionary spirit to the advising of our monograph series, which focuses on exploiting those kinds of high-resolution observations he strove to facilitate: stellar as well as solar, from ground and from space, led by some whom he had trained, inspired, and habituated to being disputing collaborators.

We have been stimulated and grateful—for his M-star initiative, for his series advice and support, and for his collaboration in using the series material to document the need for further spatial programs such as an IUE-clone devoted to clarifying stellar aerodynamic mass loss by studying its variability. After 47 years of such training, counsel, and interaction, I am desolate at its end.

The discovery of Leo's cancer forced him to drop his commitment to the M-star volume and to curtail other activities. Fortunately for the series, he had already completed his article on mass loss and strongly shaped the volume's evolution. Our expectation and hope that he would eventually best this illness (as he had so often reversed an initial setback in his programs in science and the organization of science) precluded that we dedicate the volume to him while that hope for life and rejuvenation remained for him and Bd. Now we can only star his name on the title pages of the remaining volumes, emphasizing that role of Senior Advisor to the series, in all respects, which he filled so completely.

Dick Thomas

PERSPECTIVE

This volume was first organized in the early days of the monograph planning; it then was reorganized when we realized how much the IUE was changing accepted ideas on stellar atmospheric modeling. During this reorganization, Stuart Jordan was highly preoccupied with trying to "save" another of NASA's pioneering satellites—the ill-fated Solar Optical Telescope. In consequence, he left much of the reorganization and planning of this volume to me while unselfishly carrying the burden of maintaining NASA (Goddard Space Flight Center and Headquarters) support for these monographs.

Planning such volumes is like programming your own tutorial on subjects you would like to know much more about, but have not had the opportunity to get the information through your own detailed work. Therefore, I am deeply grateful to those who have joined to tutor the readers and myself. With most of them, I have had the pleasure of interacting over many years. I have learned from, and disputed with, Anne Underhill for some 40 years. I disagree with much of what she presents here, which is all the more reason to give an essentially free hand to one of the stalwarts of hot-star studies—one who in later years has discovered and vigorously propounded the virtues of the solar/stellar connection.

When a small group of us founded JILA in 1962, we focused on laboratory and nonequilibrium thermodynamic astrophysics and on solar work. We later realized our glaring omission of stellar observational programs and facilities, which should have paralleled those in solar studies exemplified by those pursued at Sac Peak, which had long guided us. We asked Peter Conti to correct this oversight by joining us and taking such program in charge. In the reorganization of this Wolf-Rayet and O-Star volume, we did the same. His and Anne Underhill's current viewpoints differ on many features of those hot stars; the reader will get a superb overview by reading them in parallel. These two astronomers are thus the two editors of this volume.

We have already dedicated the monograph series to Daniel Chalonge, together with Cecilia Payne-Gaposchkin. Chalonge's right arm of many years, Lucian Divan, kindly joined us on this volume, working with Anne Underhill. Marie Louise Burnichon-Prevost joined those efforts as a joint product of the Chalonge and Conti schools.

Kathy Garmany, Peter's long-time collaborator, continues with him on this volume. As remarked later in this Perspective, Conti has reflected the growing emphasis on variability of hot stars generally—initiated in the series by Vera Doazan's focus on it for Be stars in the B-Be volume—by bringing to us two of the new generation of careful observers who also focus on variability of the OB stars broadly, and so overlap the Be work. Dietrich Baade, a pioneer in the study of nonradial pulsational phenomena, is a product of Mme. Seiter's active group in Muenster. (She and her husband, Helmar Duerbeck, are involved in Margherita Hack's cataclysmic-star volume; so the interrelation among the volumes continues.) Huib Henrichs enters as one of the most vigorous explorers of that variability delineated by the transient presence of the narrow high-velocity far-UV line components, the most extensive observations of which were in the Be stars.

Finally, and reflecting the most recent of the "continuous" reorganizations of this volume—to reflect new advances in observations and thinking—we are grateful to one of Unsold's "grandchildren," and his own offspring: Rolf Kudritzki, plus Rolf's students, J. Puls and A. Pauldrach. In the mid-1960s, Kurt Hunger, most-recently Unsold's successor at Kiel, came to the JILA Visiting Fellow's program to continue his He-star studies in a non-LTE "atmosphere." He interacted extensively with all of us there; returned to build a group in West Berlin, where he recruited Kudritzki; then returned with his entire group to Kiel. From there, Kudritzki was invited to reinvigorate the Institute for Astronomy in Munich, which he is now doing. I disagree with many details of his current mass-outflow elaboration and extension of what Castor, Abbott, and Klein did in JILA, but his work has produced a badly needed extension of the solar/stellar connection into atmospheres which are greatly extended, opaque in the far-UV, and produced by high-velocity mass outflow.

I trust the reader will not find this volume oriented toward my own prejudice as to how things should progress in clarifying the relation of stellar atmospheres to stars as open thermodynamic systems; as mentioned, it is a tutorial for those of us who would like to learn, not dictate.

Now, let me try to put the above efforts into the perspective of the series.

I. ONE ORIENTATION ON THE SOLAR/STELLAR CONNECTION

At the epoch when this monograph series was organized, just as the International Ultraviolet Explorer (IUE) was being launched, popular wisdom still excluded nonradiative energy fluxes and local energy dissipation by them from playing significant roles in the atmospheric structure of hot stars. Certainly the OB, probably the A, and the Wolf-Rayet were considered to be too peculiar for modeling. This orientation differed completely from that introduced by the preceding three decades of solar studies, which strongly disagreed with the basic thermodynamics underlying standard classical theory and modeling of stellar atmospheres. They rest on phenomena which appeared to be associated with nonradiative heating of the upper solar atmosphere: (1) microscopically—a rise in T_e above the predictions of standard stellar-atmospheric theory, and (2) macroscopically—a mass loss, excluded by standard stellar-structural theory, via aerodynamic outflow of the hot corona. The major problem was to decide among some combination of three alternatives: (a) nonradiative energy flux not transporting mass causes both of these phenomena simply by its energy dissipation, (b) the mass outflow itself produces the nonradiative dissipation, or (c) these two nonradiative fluxes are independent but associated in origin, size, and effect on atmospheric structure. Because neither of these nonradiative fluxes, hence none of the three alternatives, were predicted by classical theory, they all required its revision by following these observationally delineated guidelines, not speculation. The revision is not minor thermodynamically; it demands that the Sun be an *open*, not *closed* as imposed by classical theory, thermodynamic system. The origin of the mass loss defining the openness and its relation to stellar structure—atmospheric and internal—become fundamental problems in formulating a self-consistent thermodynamics for modeling stellar structure and evolution.

The primary thermodynamic question in understanding and remodeling stellar structural theory—atmospheric and subatmospheric—was whether such profound changes from classical speculative theory as demanded by these high-resolution solar observations are peculiar to it and similar stars or apply to all stars. The corollary question was whether other stars, when eventually observed under equally high resolution across a wide spectral range, will require additional thermodynamic changes, not present or obvious in the Sun, to understand and model such data. The *symbiotic* solar/stellar approach to the study of stellar structure rests on investigating these two questions: *empirically*-theoretically.

We organized this monograph series around this approach; we cannot impose it on all authors any more than they can impose classical theory on all stars. We can try to put their contribution to the volumes into its perspective.

A nonradiative atmospheric heating, with its origin in a nonradiative flux, was not the first observed phenomenon which cast doubt on the thermodynamic adequacy of classical stellar structural theory, as applied to the Sun. It was simply the first identified cause of one aspect of solar "peculiarity" (relative to standard theory) that was made quantitative. Other aspects were long familiar: anomalous atmospheric extent had been known for a century; symbiotic spectral features—simultaneous high and low ionization—had been known for half a century; and the existence of a 1 to 10.10^6 K corona had been known for several decades. All these observed atmospheric features had introduced doubt that the classical speculative quasi-thermal theory of stellar structure could represent the atmosphere of the Sun. If so, this theory is inadequate for both atmosphere and whatever subatmospheric nonthermal modes produce the nonthermal fluxes perturbing the atmosphere. There exist similar anomalous features in other stars, including some types of stars called "peculiar" because such features are so strong in their visual spectrum. The Sun, where these features are weak, is likewise "peculiar" because its proximity makes weak features observable. Therefore, the same doubts have existed, historically, among Sun and other peculiar stars as to the adequacy of classical theory to represent their actual structure—of both the observed atmosphere and the inferred interior. We have emphasized, rather than ignored, peculiar stars in organizing these monographs, and continue such emphasis in our own attempts to understand the thermodynamics of stellar structure.

The demonstration of the pervasiveness of nonradiative heating throughout much of the solar atmosphere was uniquely instructive for several reasons. First, it was those high-resolution observations of the height variation of the solar spectrum during eclipse, made possible by the combination of solar proximity and the existence of the Moon, which showed that at least one of these features of "peculiarity"—an anomalous rise in T_e—begins quite deep in the solar atmosphere, near τ (visual-continuum) ~ 0.001. Historically, this region, where classically the centers of many strong lines are formed, was pictured as a *cold* reversing layer. Second, these high-resolution solar observations also facilitated the parallel development of that thermal nonEquilibrium thermodynamics necessary to diagnose these high-quality observations, *if* their results were not to be prejudged by imposing that quasi-Equilibrium theory which they put into question. The breadth of these high-resolution data permitted these new diagnostics to be: (1) verified for consistency, (2) applied to show that the outward rise in T_e is not simply a non-LTE population effect, but requires a nonradiative heating that begins deep in the atmosphere, and (3) used to demonstrate both theoretically and empirically that such non-LTE (local thermodynamic equilibrium) effects could produce, even in regions of rising T_e, stronger absorption lines than by classical theory. A stronger absorption spectrum need not imply a stronger T_e decrease; it can be quite compatible with a T_e-rise. The centers of all the strong solar absorption lines are formed in a heated chromosphere, not in a cooled upper photosphere. All this contradicted classical theory.

Overall, therefore, these solar high-resolution quantitative measures of T_e (atmospheric height) in the low chromosphere produced a picture differing from both: the classical model of no nonradiative heating; the earliest non-RE (radiative equilibrium) model consisting only of a superhot corona. The actual nonclassical solar atmosphere begins with a very extended region where $T_e \sim T_{eff} \pm 1\text{-}2 \cdot 10^3$ K: initially, as T_e follows a non-LTE cooling, much flatter than an LTE one; then, an extended region of slow non-RE rise in T_e, but demanding a nonradiative energy input 100 times that needed to maintain a 10^6 K corona. Only when that opacity source which dominates the radiation transfer in the continuum—the hydrogen LyC in these atmospheric regions—departs from being optically thick does T_e rise abruptly, from some 8000 to some 20000 to 50000 K. It then exhibits several short plateaus, corresponding to the H-Lyα and nonhydrogen "impurities" controlling the radiative loss, before rising rapidly to $1\text{–}10 \cdot 10^6$.

Far-UV studies confirmed the essential features of this model of the nonradiatively heated subthermic outflow regions of the solar exophotosphere. We see how essential are detailed measurements of the actual $T_e(r)$ and its comparison to RE models at the beginning of any anomalous atmospheric region, before drawing conclusions on the presence or absence of that nonradiative heating excluded by classical theory. We particularly need such studies for nonsolar-type stars.

In essence, these solar attempts to make quantitative one of the long-known *peculiarity* aspects demonstrated something more important then the particular result. *"Peculiarity" is relative to some atmospheric-structural theory constructed from preconception as to what stars should be to satisfy the historic picture of them as unchanging, except on evolutionary time scales.* Evolutionary time scales are estimated relative to an internal-structural modeling which is adopted to conform to the most "quiet" thermodynamics known—linear nonEquilibrium; locally, Equilibrium distribution functions for microscopic energy states are imposed, even if one allows small, very subthermic motions. One admits gradients in Equilibrium-state parameters, but restricts them to be small enough to satisfy the linearity conditions. However, a star satisfying such linear nonEquilibrium cannot allow nonlinear conditions to occur at its outer boundaries because they will propagate downward into nonboundary regions unless strong damping mechanisms are shown to be present. And *all* the empirical solar/stellar phenomena labeled "peculiarity" have this nonlinear character. Once the diagnosis of an observed T_e rise to the boundary was shown to demand a nonlinear spectroscopic diagnostics and atmospheric structure—nonlinear radiation transport in optically thick media; and to produce a very large T_e rise, exceeding values permitted by linear boundary conditions; then, the linearity imposed by atmospheric and internal structural models became invalid. Attempts to "patch up" linear modeling, rather than simply investigating a less restrictive thermodynamics, become futile.

Therefore, the change in thermodynamic character demanded by solar observations—over speculative theory that the stars are *linear* (no non-LTE effects arising from the presence of a boundary), *quasi-thermal* (no superthermal velocities), *closed* (no mass loss), and *time-independent* (in times short compared with evolutionary) thermodynamic systems—is basic. Some stars, at least the Sun and those like it, are *nonlinear, nonthermal, nonEquilibrium, open* thermodynamic systems—*possibly variably so*, in times that are short compared with evolutionary. One had best turn immediately to observations for guidance on what thermodynamics is demanded by which stars, not by speculative theoreticians.

At the epoch cited, hot-star theoreticians believed that there was no evidence, either observational or theoretical, which demanded that *all* hot stars depart from the standard classical model for such stars. As regards any universal departure from strict classical modeling, they were ready to accept only an adaptation to radiative equilibrium photospheres of that thermal non-LTE diagnostics and atmospheric modeling which we had developed for the solar non-RE exophotosphere.

For some—but not all—varieties of hot stars, evidence of some kind of mass outflow was generally accepted, to which we return below. It was therefore acknowledged that these particular hot stars needed nonclassical modeling of then undefined varieties. However, there was no general recognition of a broad nonclassical solar/stellar thermodynamic similarity arising from the presence of a common set of fluxes other than the radiative. Actually, astronomers more generally emphasized a solar/stellar *dis*similarity in the occurrence and effect of nonradiative energy fluxes, especially relative to mass outflow. At that epoch, a nonradiative heating, and subsequent thermal expansion, of the corona were still considered to be the origin of the solar wind. In *none* of those hot-star varieties showing mass outflow did theoretical modelers admit a role for nonradiative energy dissipation and atmospheric heating.

We note two additional aspects: (1) in the solar case, nonradiative energy dissipation is generally ascribed to fluxes that do not transport mass; (2) any star having a mass outflow has a nonradiative energy flux—the question is whether it is dissipative. Those of us who have spent time in aerodynamics laboratories realize that variable superthermic mass flow, especially involving transitions between sub-, trans-, and superthermic regimes is *usually* dissipative. *A priori*, therefore, one cannot justify a contrary situation simply by hypothesis. Consequently, at the epoch cited, one inquired as to the ubiquity of superthermic mass outflow in parallel to the existence of nonradiative energy dissipation and sought any empirical association between the two. Again, the Section I, paragraph 1, solar choice, but broader. Can there exist open thermodynamic systems of the stellar type (i.e., can we have objects with a thermal radiative energy flux and a nonthermal mass-flux) without the system also producing a dissipative nonradiative energy flux? If so, are mass flux and radiative fluxes associated or independent? If not, are all fluxes associated only in the sense of nonthermal modes in the star necessarily existing to produce all these independent fluxes, or are some produced by the nonthermal modes and then produce others? The Sun gave one kind of answer, based on extensive high-resolution observations; hot-star theoreticians gave another kind of answer, based jointly on lack of theory predicting nonradiative heating and on apparent lack of observations contradicting its presumed absence.

Thus, one should keep in very clear perspective those basic thermodynamic changes demanded by solar studies, their observational foundations, and solar/stellar thermodynamic similarities wherever they may arise. Stars are unique concentrations of matter and energy in nonEquilibrium environments. Determining why and how they produce what kinds of fluxes of matter and energy back to that parental medium, especially in contrast to speculative theories that limit such fluxes to radiative, should tell us much on stellar structure and evolution. Whether these fluxes are *driven out* from stars by a variety of nonthermal storage modes/structures of stars, or whether they simply reflect *nonlinear thermal evaporation* from the outer layers of a thermal storage object, is a major question. All these problems require resolution by empirical/theoretical study, not by conjecture. This is the orientation of our monograph series, and of our attempts to put any solar/hot-star thermodynamic similarities and differences into clear perspective regarding the solar/stellar symbiotic connection.

II. ONE PERSPECTIVE ON THE HOT-STAR SOLAR/STELLAR CONNECTION GAINED BY STUDYING SOME HISTORICALLY PECULIAR HOT STARS RELATIVE TO THE SUN

A. Quasi-Steady-State Configurations: The Wolf-Rayet Stars + Sun

Solar observations in the visual spectrum, properly diagnosed under a nonEquilibrium thermodynamics developed for this solar application, showed: (1) above some atmospheric level, a monotonic outward rise in T_e, produced by nonradiative heating and departing from radiative equilibrium already at particle concentrations near 10^{15}; (2) an accelerated mass outflow of unknown origin which reaches the local thermal velocity, invalidating hydrostatic equilibrium, only at radii $\gtrsim 2\,R_\odot$, where T_e has increased to $>1.10^6$ K and particle concentration has fallen to 10^5–10^6. The actual solar mass loss is small: $\sim 3.10^{-14}\,M_\odot$/yr. Because of this, the highly superthermic maximum outflow velocities, 400 to 800 km/s, have thus far been measured only by particle collectors near the Earth, not spectroscopically in the usual solar atmosphere, the only source of information on "usual" stars; (3) a *symbiotic/composite* spectrum, with energy range from Fe I lines to forbidden lines of Fe XIV, hence originating in an ensemble of thermodynamically differing atmospheric regions; (4) a variety of phenomena, usually associated with magnetic fields, showing variability—in times *much* shorter than evolutionary—in both radial and nonradial distribution. These four aspects are usually called the "fine-structure," with a wide range in spatial and temporal scales. For the moment, we concentrate on features

(1) through (3), for which, in the Sun, we have only ambiguous evidence on variability. The surface-integrated mass outflow shows minor variability; the maximum outflow velocity varies strongly by a factor of 2. Whether this ambiguity can be clarified by detailed study of the phenomena of type (4) remains a question. By focusing on phenomena (1) through (3), we seek to put into perspective the major discords with classical theoretical models, which also focus on spherically symmetric atmospheres, in their hot-star context.

The solar far-UV adds to its visual spectra by showing highly ionized species, including resonance lines from ions whose subordinate lines are observed in Wolf-Rayet, but not solar, visual spectra. These same resonance lines appear in the Wolf-Rayet far-UV. Therefore, by contrast to the large difference in maximum ionization between hot and cold stars classically predicted and shown in their photospheric spectra, the hot Wolf-Rayet stars and cool Sun show a more similar range in ionization when the combined far-UV and visual spectra of each, produced by the entire (classically anomalous) set of atmospheric regions, not only by their photospheric spectra, are compared. If anything, because of the forbidden coronal lines observed in the solar visual spectrum, the Sun shows the greater ionization range. Because O VI exists in some Wolf-Rayet *visual* spectra, a higher ionization level than classical photospheric models predicted for the hottest stars considered, however, some of us conjectured that Wolf-Rayet stars are objects in which *only* nonradiatively heated atmospheric regions, no RE and hydrostatic equilibrium (HE) photospheres, were observed. That the Wolf-Rayet spectrum is essentially an emission-line one supported the idea. But the relative contribution to emission-line production coming from a line-formation region: that is, either much hotter or much more extended than the region of continuum formation was unresolved. Because the continuum was observationally ill-defined in the visual, and unobserved near its expected maximum in the far-UV or XUV, the picture was unverifiable.

Therefore, at least the solar atmosphere—which, by classical theory, is energetically homogeneous, with that energy level fixed wholly by the solar radiative energy—is actually observed to be multiregional, with the energy level of each of the exophotospheric regions fixed by the interaction of a variety of fluxes, and much exceeding the classical. The composite spectrum resulting from this multiregional structure resembles much more the line spectral variety, and the energy level, of the century-long peculiar Wolf-Rayet stars than do those classical theoretical models of the Wolf-Rayet stars' hot-star cousins. The question at that epoch was whether there exists the same kind of incompatibility between actual star, and classical theoretical, atmospheres for *normal* hot stars. At our present epoch, this volume's authors give an ambiguous "yes" relative to "the same." In this Perspective, I aim for perspective on this ambiguity via the peculiar stars.

The preceding remarks prove nothing on the structure of the Wolf-Rayet photosphere—the region of τ (continuum) ~ 1—nor do they demonstrate that its exophotospheric structural pattern resembles the solar. Those Wolf-Rayet observations existing at the considered epoch demanded only an atmosphere energetic enough, dense enough, and expanding rapidly enough to produce an ionization up to at least O VI in regions expanding at spectroscopically observable velocities of some 2000 km/s, with any hotter regions being too small or too dense to produce detectable forbidden coronal lines. This is in contrast to a solar atmosphere producing O VI, Fe XIV, etc., but at highly subthermic outflow velocities and low densities, reaching 800 km/s outflow only at densities too low for spectroscopic detection. Particle concentrations there are 10^3–10^2, compared with 10^{14} to 10^{12} for regions of superthermic outflow in the Wolf-Rayet stars. Therefore, although comparable thermal ionization levels, comparable macroscopic outflow velocities, and somewhat different but similarly high-ionization emission-line spectra exist in both the Sun and Wolf-Rayet stars, we did not know, empirically or theoretically, the relative contributions of the several nonradiative fluxes in producing them. At that epoch, the Sun/Wolf-Rayet comparison only suggested that those changes in the thermodynamics of classical modeling demanded by the Sun were neither restricted to it and similar cool stars, nor excluded from applicability to at least *some* hot stars. The question is how many and which.

As background for considering hot stars that are less peculiar than the Wolf-Rayet stars, we can expand several of the preceding comments by being more explicit on why some of these solar spectral features have their observed characteristics. One sees visual spectral evidence for its 10^6 K corona in the solar case—the first-found direct evidence for nonradiative heating—*only* because densities are sufficiently low in this solar region to permit formation of forbidden lines. A second condition on coronal visibility in such visual spectral forbidden lines is the corona's occupying sufficient volume to produce earth-detectible forbidden-line emission at solar (not stellar) distances.

Therefore, a star producing a high density or a too small volume (even a solar size) corona would not signal its presence in its visual spectrum observed at the Earth. Because O VI—an ion formed in the transition region between solar corona and chromosphere—is the highest ionization species with resonance lines visible in the far-UV, such a star would provide no evidence of a 10^6 K corona in even its combined visual plus far-UV spectrum. At most, we could observationally infer the existence of regions at whatever T_e are necessary to produce O VI, *if* we correctly understand the conditions that produce O VI in such strongly nonEquilibrium exophotospheres. To infer the existence of 10^6 K regions requires XUV and X-ray observations, not then existing. They exist at today's epoch, and show energies in the solar coronal range, with much larger intensities. We will return to this point.

Why such a low-density solar corona? Simply because the mass flux is so small. A constant mass outflow below a certain size does not perturb an HE photospheric density distribution. So such a mass outflow, of given size, has an outflow velocity at each height—corresponding to the photospheric density there—which is specified directly by the HE photospheric model. Physically and observationally, it is useful to define that the photosphere ends when nonradiative heating produces a chromosphere, but the set of HE regions ends only when the mass-outflow velocity has accelerated to near-thermal values. Until just below that thermal point, the density gradient is effectively exponential; the density drops rapidly. Above the thermal point, the density gradient is $\sim(\text{velocity})^{-1}$, a much smaller gradient. Therefore, the larger the mass outflow: the larger the outflow velocity at each height in a given HE atmosphere; the lower the height of, and the larger the density at, the thermal point; and the larger the density throughout the overlying atmospheric regions. The small solar mass flux gives a small (indeed observationally negligible) photospheric outflow velocity, a long atmospheric region of subthermic outflow, a long region of exponential density gradient, and therefore a very low-density corona. The Wolf-Rayet stars, with a mass outflow 10^{10} larger than the Sun, should have very high-density exophotospheres—and high density chromospheres/coronae *if* a nonradiative heating to produce them exists.

A simple calculation of where an HE photosphere should no more exist, as a function of mass-outflow size, is illustrative. Stars do not differ drastically in the photospheric densities at optical depth near 1. The Sun, a cool dwarf with log g near 4, has a particle concentration of 10^{17} there; a hot star with log g of 4 has 10^{15}. However, if each star has an extended atmosphere resulting from some combination of nonradiative heating and mass outflow (as for the Sun), they can differ very dramatically in densities at the thermal point, and in the overlying atmospheric regions, as we saw previously: a factor 10^{10} between Sun and Wolf-Rayet star. We ask what is the largest mass outflow a star could produce and still have an HE photosphere at τ (continuum) = 1. Mass outflow = $4\pi R^2 \rho U$. We set $U = q = (kT_e/\mu)^{1/2}$; non-HE effects begin when $U \gtrsim q/3$. For a hot star with $R \sim 10 R_\odot$, we obtain mass outflow $\lesssim 10^{-4} M_\odot$/year, which is essentially the observed value for Wolf-Rayet stars. Possibly we should not be surprised to find no indication of a classical photosphere—RE and HE—for Wolf-Rayet stars. At $\tau \sim 1$, an atmosphere with mass outflow exceeding this limiting value is already expanding superthermally; it has traversed the transthermal region where aerodynamic heating is customary. If this is indeed the thermodynamic configuration of Wolf-Rayet stars, we had no basis for comparing standard atmosphere (RE,HE) photospheric models with observed spectrum to infer a T_e-excess, hence heating. We could only say that, because the line spectrum is wholly an emission one; because ascribing this wholly to an extended atmosphere at τ (cont) ~ 1 is difficult; and because the lines are strong enough that the source-sink terms must be included (a wholly scattering treatment is excluded) in any (obviously non-LTE) radiation-transfer treatment of line formation; an inference of some amount of nonradiative heating seemed necessary. We therefore suggested that, by comparing Sun and Wolf-Rayet stars, we might be comparing the extremes in size of mass outflow—one in which it does not perturb HE for several radii, and one in which HE is already excluded in the lowest observable regions—and finding some nonradiative heating at each of the extremes.

The specific source of such nonradiative heating, or the specific kind of nonradiative energy flux which is dissipative, need not be the same across this range in mass outflow. We proposed only that the thermodynamic character be the same: all would be objects with the full range of nonthermal fluxes of mass and energy returned to the parental medium. Clearly, the size of each flux need not be the same for each star. One can only ask observationally, given the present state of any theory, if having the same full range of fluxes guarantees the same set of exophotospheric regions. Even if this is so, one cannot demand the same set of values for the thermodynamic parameters specifying the particular thermodynamic state of each of the several atmospheric regions—T_e, outflow velocity, particle concentration, distribution functions for particles and photons, etc. The values of these parameters will depend on the values of the several fluxes and of the lower boundary conditions at the subatmosphere/atmosphere interface, all of which can vary strongly across the range of stars from Wolf-Rayet to Sun and beyond.

If the thermodynamic logic of the preceding is correct, we should expect to find some association between the existence of mass outflow and the existence of nonradiative heating in any star which has a mass outflow lying between the extremes spanned by the Sun and the Wolf-Rayet stars. Furthermore, if that thermodynamic character implied by the range of peculiar stars exhibiting extended atmospheres—*open* system mass loss—is general rather than particular, *all* stars will show this (mass outflow, nonradiative heating) association to some degree. The problem, as exemplified by the Sun, is detection: given that the size of the mass flow and the conditions at the lower boundary (i.e., the HE,RE region) must be permitted to differ between stars. The amount of heating fixes the extent of the non-RE but HE region. The size of the mass outflow fixes the beginning of the non-HE region, hence the density in the several atmospheric regions and their spectroscopic detectability. *Any existing mass outflow and nonradiative heating may or may not be associated in origin, but their effects on atmospheric diagnostics are interlinked.*

Thus, in exploring a possible extension of the solar/Wolf-Rayet "similarity," a first focus lies simply on the search for the existence of *associations*; one must ask *whether* they exist, *not why*. Questions of origin and cause are deferred—cf. the frontispiece of each volume in the series. We had still the solar, paragraph 1 of this chapter, ambiguity between origins and associations of the nonthermal fluxes and their relation to subatmospheric structure. The observational and diagnostic basis existing at the series-organization epoch needed expansion; any general modification of theory would follow, not lead. The parameters of "association" were the presence of those spectral lines lying intermediate between photospheric and solar/coronal ionization levels, possibly with XUV and X-ray additions, on one side; and first superthermic velocities (measuring nonlinearity), then superescape (measuring openness), on the other side. Microscopic, macroscopic linkage.

Such supplementary information came in three stages. Rocket observations during the decade preceding IUE—and the organization of these monographs—had established a close link between the Wolf-Rayet stars and some hot supergiants in that historic Wolf-Rayet characteristic of superthermic, sometimes superescape, mass outflow. Moreover, those far-UV spectral lines which showed mass loss in Wolf-Rayet stars, and nonthermal heating in the Sun, were those which showed the superescape mass outflow in the OB-sg. Even a few main-sequence, so less luminous and less easily observed, hot stars showed such superthermic, possibly superescape, mass outflow—but of much smaller size, hence with spectroscopically observable regions being closer to the star—via diagnosis of the same spectral lines.

IUE observations have provided the second stage of new observations. Since its launching—and that of the organization of these monographs—the variety of stars—both hot and cold, sg, g, MS—which show mass outflow of a range of velocity, and in a variety of spectral lines, has continually increased. As summarized by the authors of this volume, the O VI, N V, C IV, and Si IV resonance lines remain prominent diagnostic indicators of such superthermic and superescape velocities among the O stars. As summarized in the B-star volume, the number of B main-sequence stars that unambiguously exhibit such outflow via such spectral lines decreases toward the cooler stars: O VI is observed in only the hottest, then C IV, etc.; and maximum inferred velocities decrease similarly. The most prominent "peculiar" B stars—the Be, to which we return below—show these effects more strongly than the B-"normal," extending to cooler, less luminous stars. Some among the A stars—particularly the Ae stars—also show these phenomena, with lower velocities: not always superescape, but always superthermal relative to photospheric T_e. Therefore the "associated presence" of the two phenomena—mass loss, and ions whose presence requires nonradiative energy dissipation to raise T_e in at least *some* stars whose photospheres are cooler than some minimal T_e—has become well established. What produces the mass loss and nonradiative dissipation, and any relation between the two in the previously stated sense, remains to be settled. The authors of this volume address these questions from a variety of viewpoints.

One has a mixture of two possible effects, in diagnosing the *apparent* decrease in velocity amplitude, and in observed strength/presence of the O VI—Si IV lines, with decreasing luminosity. There can be a real decrease, and there can be a part simply representing a detectability limitation, which arises from those effects of decreasing particle concentration in the diagnosed atmospheric region which we have already discussed. Figure 3-33 in Thomas, 1983 is a very illustrative diagram. It assembles all measures of the maximum velocity from stars across the HR diagram that were available in 1983. When measures at different epochs gave different velocities for a given star, both were plotted and are joined by a line. (We will return to this variability.) It is interesting to note that, if the solar measures are excluded, no star cooler than B9 shows an outflow velocity $\geqslant$500 km/s—from wholly spectroscopic measurements. When particle-collector measurements of the solar outflow are included, the Sun exceeds this "observational limit" by a factor of 2. We ask what particle-collector measurements of star samples across the HR diagram would produce.

The resolution of these problems is fundamental for establishing the basic thermodynamics of stellar modeling, especially its broad uniformity or not, as already discussed. Before further comment, we add two kinds of additional information produced by spatial observations.

Observations in the X-ray spectral region from several space observatories show coronal-level thermal X-rays in some stars all across the HR diagram. For any given (visual) spectral class, the scatter in X-ray luminosity is very large, essentially two orders of magnitude. There is also a general decrease from the hottest to coolest stars—a factor of 10^4 between O and M stars; and an ill-defined but definite excess in sg and g over MS. Nonetheless, the implication is very clear. In those stars showing such X-rays, there exist some atmospheric regions where solar-coronal level T_e occurs. The regions must be large enough in each such star to produce sufficient X-rays for detection. Whether the difference in X-ray luminosity between two stars of the same (visual) spectral class arises from difference in T_e, in volume of the "hot" regions, or in presence of overlying cool absorbing regions is not clear. (But see the comments below on Be stars.) It is clear, however, that nonradiative energy dissipation exists, to varying degree, in some atmospheric regions of hot stars, as well as warm, cool, and cold stars. Whether such regions are part of an atmospheric sequence resembling the solar, are part of some other kind of sequence, or lie in no sequence at all remains to be studied and decided *empirically*.

A second addition to empirical information lies in a compilation of mass-loss inferences across the HR diagram. Above, we focused on searching association between mass loss and nonradiative heating. If we consider the variation of those values of mass loss inferred from observed outflow velocity combined with some estimates of particle density in the flow region studied (hence highly model dependent), we find a strong decrease in mass loss with decreasing stellar luminosity. For the O stars, the radius differences between sg, g, and MS are not so striking that one can attribute a factor of 10^3 to 10^5 difference in total mass loss to a constant mass flux simply coming from a larger surface. The situation is more ambiguous for G-sg and G-dwarf stars. It is interesting in comparing B-sg to B dwarf, especially to Be. Among the O stars, however, the empirical strong variation of mass loss—or of mass flux—with luminosity—or with radiative flux—is striking. The trend of decrease in mass loss with luminosity continues among the B-dwarf stars. For this reason, most astronomers are convinced that the origin of mass loss in hot stars, at least in Wolf-Rayet and O stars, lies in the radiative flux. Currently, theoreticians insist on an *origin* only among the O stars, demanding only a radiative acceleration of an otherwise originating mass loss for cooler hot stars. For such theoreticians, therefore, the O + Wolf-Rayet stars remain a breed apart, in regard to *why* classical standard modeling fails, but the presence of a nonradiative heating, somewhere in the atmosphere, can no longer be ignored. As yet, its presence is not incorporated into their modeling: no aerodynamic dissipative effects.

So we have, for empirical-theoretical guidance: (1) the above emphasized association between the *presence* of nonthermal mass outflow and of nonradiative heating for stars where O VI et al. and X-rays imply heating; (2) an association between sizes of luminosity and (a) X-ray luminosity for all stars and (b) mass loss for hot stars; (3) a decrease in outflow velocity and strength of O VI, C IV, et al. toward the cooler hot stars under spectroscopic diagnostics, but "recovery" of high velocities under solar particle-collector diagnostics. We therefore have the same dilemma between causality and associated-origin for a number of energy and mass fluxes, as we did for the strictly solar case (Section I, paragraph 1) when we diagnose the observations under the assumption of time-independence of all fluxes and atmospheric structure.

The authors of this volume add to the observational variety, and discuss these dilemmas/problems, from various viewpoints. Some restrict themselves mainly to compiling and discussing the observations. Others seek an origin of nonradiative heating via solar analogy, but focused wholly on hydromagnetic effects which, in today's solar thinking, are usually considered unimportant for such heating in those lower solar regions where nonradiative heating begins. Still others seek ways to avoid the need for such nonradiative heating in modeling the mass outflow by trying to show that the O VI, C IV, etc., lines can, for the hottest stars with largest mass outflow, be modeled under RE. In such an approach, they must produce the X-rays in regions other than those affecting the character of the mass outflow.

To conclude this section on Sun/Wolf-Rayet symbiotic perspective on hot-star studies, we focus on the physics underlying the just-mentioned discussion of O VI and RE: we abstract a solar result on the production of "superexcitation" under conditions of a small T_e rise combined with large local opacity and strong non-LTE.

At the beginning of this Perspective, we stressed the necessity for very precise determination of $T_e(r)$ in any RE-to-non-RE transition. Empirically, its departure from T_{eff} is systematic but small, over most of the solar (upper-photosphere to lower-chromosphere) transition—although the latter is the locale of strongest nonradiative heating; and ends in a rapid rise. A second major empirical result of those early solar studies was a superexcitation of hydrogen: a strong overpopulation of its first excited level, $n = 2$, relative to classical models, which was needed to give that Balmer opacity demanded by observations. Once the small T_e rise was combined with the large opacity increase demanded by a self-consistent non-LTE, the origin of the superexcitation became clear.

Iteration from the thin-atmosphere approximation of collisional ionization and radiative recombination produces a locally opaque LyC—and Ly-lines—in the low chromosphere. This gives the non-LTE factors $b_1 = b_2$; hence an n_2/n_1 fixed only by the local T_e. This same result would hold if collisions dominated the rate processes, but they do not; the result is an opacity effect. Thus, for even a small T_e rise, there is an enhanced n_2 over an RE state, because hydrogen is still neutral at these heights. This at once makes the Balmer lines stronger than classical modeling would produce, and makes photoionization from the second level the process that determines the degree of ionization of hydrogen. Effectively: hydrogen becomes an ion with 3.5-ev—rather than 13.5—ionization potential; the $n = 2$ level is the ground state of the ion; the Balmer lines are effectively resonance lines. The chromosphere is transparent to the BaC, but not to the LyC; it is the photospheric BaC which fixes the ionization balance. The local LyC is fixed by local conditions—increasing in strength with height; equaling the local $B_\nu(T_e)/b_1$. (When the local opacity in the LyC decreases, the required transfer equation in the LyC "phases out" this superexcitation effect; $b_1 > b_2$ becomes progressively fixed by the thin-atmosphere collisional effects; ionization degree becomes density-independent.) One of the first "successes" of this diagnostics/scenario was the prediction of a hydrogen LyC much stronger (a factor of $\sim 10^6$) than the classical model—according with rocket observations. Indirectly, therefore, this "superexcitation" of the $n = 2$ level of hydrogen, coming from a very small T_e rise, but a large non-LTE opacity, implied this *non-RE*-originating "excess" in the LyC.

Apparently, the demonstration by Kudritzki et al. of a strong increase in concentrations of O^{+5}, etc.—which arise from the high local opacity of the large mass outflows of the O and Wolf-Rayet stars under proper non-LTE treatment (cf. Chapter 4)— is thermodynamically similar to the above hydrogen superexcitation. They explore ad hoc T_e distributions, emphasizing their nearness to T_{eff}, by contrast to the much larger T_e required using thin-atmosphere collisional processes to produce "superionization." Unfortunately, at least in their discussions to date, their emphasis lies too greatly on implying that their clarification of the "superionization" supports their particular theory of mass outflow and a minor, if any, departure from RE for these hot stars. However, *any large* mass-outflow to produce large densities; their correct, locally-opaque *in at least* the ground-level transitions of ions otherwise-marginally-present, non-LTE treatment; and some small T_e rise;—to reduce effective ionization energy to reach OVI, etc. stages—will produce the effect. It has not yet been shown that the photospheric-radiative acceleration theory they use is unique in producing a large mass outflow, particularly in its neglect of *any* nonradiative heating and its apparent inability to produce the observed variability aspects (cf. below).

The situation is a good example of we solar-symbioticists not being sufficiently aware of the thermodynamic breadth of our solar work to have discussed its broader application before the hot-star theoreticians exhibited its utility in much hotter, nonsolar stars, ignoring the solar liaison. It is also a good example of how a solar result—T_e lying near T_{eff}, even under much heating, until the dominating continuous opacity becomes small—might be combined with similar hot-star studies, in which the continuous opacity is quite different, to further clarify this association between opacity and T_e rise under nonradiative heating. One will not exploit the range of conditions across the HR diagram to study stars generally if he does not recognize thermodynamic similarity wherever it exists.

One hopes that these investigations summarized by Kudritzki et al. will be tested by an indicator of the actual $T_e(r)$—e.g., by a continuum equivalent to the solar LyC. In these hot stars, such continua lie even farther into the XUV. One notes with interest the details of the steep rise in the XUV continuum of the hot white dwarf, HZ 43 (Malina et al., 1982). We also hope that these studies by the Kudritzki group, with their detailed non-LTE programs, can be extended to exhibit just how low a T_{eff}, under what size mass outflow, and with how much non-RE increase in T_e, can produce which of the "superionized" lines. When the theoretical models can produce reliable RE distributions of $T_e(r)$, we can better estimate just how much nonradiative heating is required for these hot stars, and try to identify what dissipative term(s) added to the mass-outflow equations, or possibly coming from a nonradiative energy flux not transporting mass, can produce the required nonradiative heating. Such studies are required for self-consistent modeling of any atmospheric structure arising from either steady or nonsteady mass outflow—*and* for subsequent inference on the subatmospheric structure producing such outflow.

B. Time Dependent Fluxes and Atmospheric Structure: The Be Stars Versus Normal B Stars and Sun as Perspective on O-Star Variability

1. Some General Remarks. In the opening paragraph of the preceding section, we mentioned the ambiguity of solar observations of mass outflow in regard to its variability. Some years ago, effectively on the basis of stellar-structural theory and visual-spectral observation, most astronomers were convinced that hot stars showed little evidence of variability except for "peculiar" hot stars: notably the Be–Oe–Ae, known for a century; the β Ceph "pulsators," known for half a century; more recently, other small-amplitude, spectrum, and luminosity variables of short-period, etc. Hence, this volume on the O and Wolf-Rayet stars, as originally planned a decade ago, largely bypassed the variability aspect, deferring to the discussion of the Be and β Ceph stars in the B-Be-star volume.

However, over the last 15 years of rocket and satellite far-UV observations, and especially of a decade's detailed time-sequential observational studies of individual stars via the IUE, the accumulated body of evidence on variability has greatly changed assessments of its importance. Also, the modeling outlook has evolved, under the impact of both theoretical and observational focuses on trying to clarify the characteristics and origin of mass loss, which becomes increasingly recognized as pervasive among stars. *The strong long-term visual-spectral variability of the cool emission envelope, which has long been known for Be–Oe–Ae stars—but which was set aside in historical modeling attempts to focus simply on the origin and consequences of steady mass outflow—reemerges as their principal distinguishing characteristic.* Far-UV observations link this cool-envelope variability to that of those underlying hot regions of the exophotosphere, whose existence the far-UV solidified. But more modest variability of even the normal O—even of the Wolf-Rayet—stars now appears to be much more significant for understanding mass outflow, its effect on atmospheric structure, and its implication on internal structure, than was earlier recognized.

At one time, the Oe and Be stars were considered to be unique among the OB stars in demanding mass outflow—to create that extended atmosphere in which arise those emission lines which define these stars. Classical standard theory was *a priori* inadequate for them because its thermodynamic basis excluded such mass loss. Now, from far-UV observations, *all* OB stars are recognized to have a mass outflow of some size, and a "revised" standard theory attributes the mass loss in normal hot stars to an outward acceleration by the photospheric radiation field. If one asks what distinguishes the Oe–Be mass outflow from that of normal stars—and what distinguishes the Oe–Be kind of extended atmosphere, characterized by low-ionization emission lines, from the extended atmosphere produced by *any* star with aerodynamic mass outflow—the observational evidence suggests it is the striking variabilities observed (and associated) in both far-UV and visual spectra and the (associated) variable existence of the cool outermost envelope. *Therefore, the revised standard theory also fails for Oe–Be stars, equally a priori: because that theory imposes time-independence and current observations show variability*; but also because that theory contains no mechanism for producing that variability in photospheric radiation field required, *if* that theory is applicable, to match the several observed aspects of Oe–Be variability. One asks if such theory can fail for Oe–Be, yet suffice for normal OB.

Even more recently, more extensive study of individual normal OB stars exhibits some aspects of mass-outflow variability similar to those of the Oe–Be stars, even if not as striking. And some OB sg are long known to exhibit some of the variable low-ionization emission lines, even if not as strikingly, which characterize the main-sequence Oe–Be. *One therefore asks if that evolution in the existence of mass loss, from being an Oe–Be peculiarity to being a normal hot-star property, will be duplicated in the evolution of the property of variability of mass loss, probably admitting a difference in degree/strikingness.*

Clearly, a major problem in forecasting the answer to that question lies in our limited knowledge of the difference between the peculiar Oe–Be and the normal OB stars in both the existence and variability of mass outflow. The IUE has been a gold mine of information in mapping the existence of mass outflow across the HR plane. It, and its urgently needed successor—which has been discussed for years, unfortunately with no concrete production/launching schedule yet resulting—*could be platinum mines, if expeditiously utilized in conjunction with ground-based facilities.* At the moment, the limited existing data, which rest heavily on such simultaneous visual plus far-UV observations, suggest that the difference between Oe–Be and normal OB stars does indeed lie in the degree of variability of mass outflow and its atmospheric consequences (cf. Section 2 which follows.) Obviously, such introductory, exploratory results should be checked in detail.

Obtaining an adequate data set to permit a penetrating study of the variability of mass loss will be a difficult challenge that must be met. Unless an ultraviolet observatory like IUE is dedicated to obtaining these data for part of its observing time over the duration of the mission, it is unlikely that they will be obtained. The IUE experience demonstrates this: the astronomical user-community for such a facility is large, and there is great difficulty in obtaining sufficiently frequent observations to build up an adequate data base for attacking any fundamentally important problem such as this one where *all* the observing time must be shared with a very large group of users representing a diffuse community of interests. Clearly, there is a need for the leadership of NASA and other space agencies to meet the challenge of setting priorities on at least part of the observing time, so that an adequate data base will result for the solution of fundamentally important problems. Equally clearly, it is our conviction that the problem of the variability of stellar mass loss is just such a problem, based on its implication for stellar evolution, for stellar subsurface structure, and for the elucidation of mass-loss mechanisms, all of which is demonstrated by the volumes in this series.

One is only bemused by studies of the evolution, galactic composition, and cosmology of objects which are *speculatively theoretically modeled/discussed as short-time invariant, but are actually real-world observed as short-time variable*, especially in that mass loss whose detailed origin is neglected in such evolutionary/cosmological studies. *Such modeling ignores the difference in physics between nonlinear open and closed thermodynamic systems*; the modeler becomes lost in his satisfaction with the mathematics of producing better computer codes for thermodynamically inconsistent mixtures of the two kinds of thermodynamic systems. Unfortunately, such computer simulation simulates neither the basic physics nor the observations. We need improvements in, not solutions of, physically incomplete equations. This requires more complete understanding of just what empirical/observational phenomena such equations must be able to represent. It is therefore absolutely critical to delineate observationally the pervasiveness of some/any kind of variability in mass outflow, and then the characteristics of this (possibly several varieties of) variability.

One notes that the production cost of an IUE clone is a very small fraction of that of the Space Telescope; the observing/diagnostics cost could be similarly less, and it could probably be launched into a 10-year orbit 3 years from go-ahead. This type of hardware, with detector upgrade, already exists. A science-oriented IUE followup, based on space results, does not.

What is needed is a dedicated mission under a dedicated program head, as opposed to a committee approach. It is quite clear as to just what observations are needed. Those few existing long-time-sequences of study of a few stars display the profound conflict between the thermodynamics of current modeling theory and actual stellar behavior. Further progress requires such studies for a variety of stars; present data suffice to put modeling and theory into question, not to map definitely their required changes. Expansion to the XUV regime would add another dimension of information, if it went down far enough in wavelength (cf. the cited Malina et al. observation as illustrative). But such XUV addition cannot replace the need for duplicating the full IUE spectral range, coupled with simultaneous observations in the visual as illustrated by Section 3 following.

2. The Summary of Hot-Star Variability in this Volume. In an attempt to at least partially address the changed outlook resulting from these newer kinds and quality of data for the O and Wolf-Rayet stars, one of the editors of this volume, Peter Conti, has invited two active workers on hot-star variability, Baade (in the visual) and Henrichs (in the far-UV), to summarize some aspects of recent progress in the subject. Because each recognizes that the strongest variability is shown in the Be–Oe stars, their summaries contain many comparisons between these and normal O stars. However, because in the different spectral regions one sees different atmospheric regions, the particular observations they discuss lead to their focusing on different phenomenological features.

Baade discusses mainly variability in line-spectral features, emphasizing that the needed simultaneous photometric observations do not yet exist for the O stars. His discussion centers on pulsations—the observational evidence for their existence and the possibility that they can act as a driving mechanism to initiate or enhance the mass outflow. Although he adopts the generally accepted viewpoint that mass loss from hot stars generally agrees well with radiative-accelerated mass-outflow theory, he emphasizes that the scatter in observationally inferred mass-loss rates at a given point on the HR diagram appears to be larger than observational error, binarity, etc., can explain. Hence, his outlook in delineating the presence and effect of pulsation lies in viewing it as a supplement—not as a replacement—to explain departures from, or scatter about, the locus of radiative-accelerated mass-loss theory. Because current pulsation theory has some ambiguity in specifying just where in the atmosphere it begins to drive an actual mass *flux*; and because there is some ambiguity whether the radiative theory initiates a particular mass *loss*, or only "filters" a particular value from a spectrum otherwise imposed (cf. Section B); the physics of a combined pulsation and radiative *origin* of mass outflow is not clear. Hence, Baade's summary is most valuable in his delineating the various aspects of *association* between line-profile variability measures of pulsation and some aspects of mass outflow.

Because, for the Oe–Be stars, visual-spectral measures cover both the low-atmosphere pulsation features and the outer-atmosphere features of the cool envelope, he also summarizes associations between these two phenomena and two atmospheric regions. In particular, for the Oe–Be stars, he attempts to delineate any associations between photospheric-pulsation indicators and V/R for the envelope emission lines. V/R is the ratio of violet to red emission peaks.

Unfortunately, the restriction to only visual-spectral features, which Baade violates only occasionally but then very usefully, means that he must ignore associations with those regions lying between photosphere and envelope. As he himself stresses, there are presently no data showing the propagation of any disturbance from the photosphere to those atmospheric regions, observed in the far-UV, where arise the variable narrow-absorption components on which Henrichs' summary focuses. However, there is now a considerable body of information on observed associations between these narrow components and phenomena in the cool envelope. Unfortunately—almost necessitated by the division of labor between the two summaries—neither Baade nor Henrichs surveys these aspects, on which information comes from simultaneous visual and far-UV observations. Section 3, which follows, gives for completeness an abstract of this material from Doazan's recent work on Be stars, supplementing that in the B–Be star volume. As stressed by Henrichs, it is necessary to begin with such Be material in the present volume as the departure point for putting variability among the O stars into perspective.

Both Baade and Henrichs stress that the presence of the variable narrow-absorption components on which Henrichs focuses seems almost ubiquitous among the O stars (cf. also Prinja and Howarth, 1986, whose statistics weigh heavily in Baade's emphasis). Henrichs' survey exhibits this evidence for the pervasive existence of variability, to some degree, throughout the O stars, normal as well as peculiar. As such, it is invaluable in showing the error of those who would consider variability as exceptional, hence ignorable, in constructing any broad theory and modeling for all but "exceptional/peculiar" stars. Instead, *from work such as the surveys of Henrichs and Baade in this volume, we must consider variability as a "normal" property of hot stars, taking the composite conclusion of this and the preceding B–Be star volumes.* The A-star volume exhibited those stars as a class in which, quoting its author, Wolff, "abnormalities are the rule rather than the exception." And among such abnormalities, or peculiarities, variability ranks high. For them, we simply have not the accumulated evidence associating mass outflow, nonradiative heating, atmospheric structure, and variability to the degree which the Be–Oe, and now normal OB stars, begin to exhibit it.

For OB stars, therefore, we continue trying to decide whether the distinction between peculiar and normal stars lies only in the degree to which these departures from some variety of "standard" theoretical models occur, or is it something more profound? If it is only degree, than clearly any "standard" modeling which does not incorporate variability in all fluxes as fundamental, with only their several amplitudes to be determined in each star, can hardly be called "standard."

Henrichs accords with Baade in trying to represent such variability as only a perturbation on, not a replacement of, a radiation-driven/originating mass outflow. Because existing theories of such outflow set aside the possibility of variability, the preceding remarks apply. For observational guidance, however, more serious is Henrichs' restriction to the far-UV, and his emphasis on the narrow-absorption components in those lines we have earlier referred to as "superionized." These components, which are variable in presence and lifetime and are usually displaced to quite high velocities, are superposed on a much broader profile, whose variability is much less. Because of its restriction to the far-UV, his summary says essentially nothing on any association in behavior between that of these narrow components and that of any visual spectral features, and this is the survey's (*a priori*) shortcoming. To see this, we abstract an overall picture of Be-star variability as it relates to mass outflow, following Doazan (1987) and Doazan and Thomas (1987, 1988), which should be read as an updated supplement by Doazan's work to the background presented by her in the B–Be volume in this series (Underhill and Doazan, 1982), as a transition to the discussion of the Be–Oe–O stars in the present volume.

3. *Abstract of Observational Evidence for Variable Mass Outflow in Be Stars.* Because the B–Be volume summarized in detail that major visual-spectral variability of Be–Oe stars—consisting of transitions between the three phases: Be (emission-line), B-shell (narrow-absorption cores in low-ionized lines), and B-normal (*visually* indistinguishable from normal B stars)—it need not be repeated here. Observationally, what has been greatly clarified is: the Be far-UV behavior, especially in association with simultaneous visual observations; and the identification of (far-UV) features distinguishing Be-stars in B-normal phases from normal B stars. In addition, the features of the (empirical-theoretical) atmospheric models proposed in that volume have been elaborated. Therefore, I focus on these aspects of variability in abstracting the cited work and references.

As discussed previously, early stars generally show variable profiles for those far-UV lines used to study mass outflow. However, all OB stars do not show the same variability behavior.

a. In "normal" OB stars, the variability is usually associated with changes in the fine structure of the broad asymmetric profiles of the highly ionized resonance lines discussed previously. These lines often show one, or several, narrow-absorption components, which seem to occur preferentially at high velocities, although low-velocity components are also observed. The far-UV variability of such stars is generally ascribed to the variability of these narrow components (cf. the chapter by Henrichs in this volume), although the underlying broad profile also varies (cf. Prinja and Howarth, 1986).

In the Be — we have as yet not as complete data for the Oe stars — the situation differs profoundly. Generally, the *entire* profile of the far-UV "superionized" resonance lines *strongly* vary in strength, velocity, and shape. When observed, variability of the narrow components is only one aspect, but an important diagnostic one, of Be variability (cf. the following).

b. Those far-UV observations presently available suggest that the occurrence/enhancement of narrow-absorption components in "normal" OB stars does not show any systematic time-dependent pattern. Such components seem to occur erratically, but it is not certain that such apparently erractic behavior reflects the real character of mass-outflow variability in these stars. It may reflect only the incompleteness of present observations for delineating any organized patterns of variability that may exist.

On the contrary, systematic and long-term investigations of a few Be stars, made simultaneously in the far-UV and in the visual, have shown that: *long-term organized variability patterns of large amplitude are exhibited in the far-UV; such patterns are associated with those, long-known, exhibited in the visual region by the Balmer emission/shell lines; and that the short-term variability features; are often associated with the long-term.* The following illustrations are taken from the results of Doazan's systematic investigations of the relations between variability patterns exhibited on various time scales; atmospheric structure; and what these imply on subatmospheric structure; for Be stars (cf. Doazan, 1986). As such, and being very cognizant of the way large-amplitude "peculiar-star" phenomena have forecast the future of smaller-amplitude "normal-star" phenomena as observational resolution improves, we should carefully ask their implications on hot-star, even general-star, modeling and theory.

For 59 Cyg (B1.5 Ve), the C IV absorption-line equivalent width shows long-term association with the equivalent width of the Hα emission line, and both with V/R of the Hα emission line.

For γ Cas (BO.5 IVe), the short-term transient occurrence of the narrow-absorption components in N V, C IV, and Si IV is not random relative to long-term variability patterns; these components appear to occur only when V/R for Hα exceeds unity.

For θ CrB (B6 Ve), the C IV resonance lines exhibit a long-term, systematic trend toward a higher expansion velocity *during a phase classified, on the basis of visual spectra alone, as "normal B."* Moreover, during the later stages of the systematic increase in outflow velocity, the C IV line often weakens very strongly; indeed, at very later stages, it is more often undetectable than visible. When it appears, however, it shows the monotonic increase in outflow velocity.

For 88 Her (B7 Ve), there appears to be evidence, coming from simultaneous far-UV and visual photometric data, which links a decrease in luminosity to an increase in mass outflow at the beginning of a B-shell phase. Moreover, the data contradict historic belief that such a luminosity drop at the beginning of a shell phase simply reflects the growing presence of the cool envelope. These present data show the luminosity drop to *precede* the appearance of the envelope and to imply a significant drop in T_{eff} by some 1000 K. Indeed, as the envelope develops, the luminosity *increases. Such diagnostics are only possible if we have simultaneous coverage of visual and far-UV spectral ranges.*

As stressed earlier, the classical peculiarity of Be and Oe stars relative to classical standard models was a mass outflow inferred from visual-spectral observations alone. "Theory" concentrated on explaining its origin, rather than clarifying the character of an atmosphere that resembles a planetary nebula more than the thin classical atmosphere. It postulated an equatorial ejection from rotational instability; hence, an equatorial-disk envelope. Visual observations alone led to outflow velocities $\lesssim 200$

km/s; modern far-UV observations give 200 to 2000 km/s, sometimes with this range in the same star at differing epochs. Some *inertial* modeling, trying to preserve the old while incorporating the new, postulates:

(1) A high-mass, low-velocity, equatorial outflow, whose characteristics are those of the classical visual spectrum, and which possibly originates in some equatorial-focused, subatmospheric "activity," such as a nonradial pulsation in modes concentrated to the equator.

(2) A low-mass, high-velocity polar outflow—whose characteristics are those of the modern far-UV spectrum, and which possibly originates in a way similar to the "normal" mass outflow of "normal" hot stars—currently generally believed to be that of the radiative-acceleration theory.

The striking, unsuppressible problems with all aspects of such models are:

(1) The variability in all Be and Oe phenomena, consisting of: (a) the classical, long-time scale, alternation among the three phases of Be stars—Be emission, B shell, and B normal, and (b) the short- and long-term variability abstracted above.

(2) The above-abstracted association between visual and far-UV aspects of variability, which imply dynamic interaction between the flow regions producing the far-UV and visual spectra. Thus far, no one has managed to suggest a way for any physically separated polar and equatorial outflows, as abstracted above, to produce such associated behavior.

(3) The fact that the high-velocity, hot, spectral lines appear in absorption—there are *no* emission components of these "superionized" lines in the far-UV of Be stars—while the low-velocity, cool, lines appear in emission, places the cool envelope outside the hot regions.

(4) The velocities inferred from the visual-spectral lines are much less than those inferred from the far-UV lines; which, given (3) above, implies decelerated outflow *above* the nonradiatively heated regions.

(5) The far-UV velocity variability suggests both: (a) a variability in the initial outflow velocity—implying a variability in the mass outflow imposed in the lowest atmospheric levels; and (b) a variability in its subsequent acceleration. The kind of variability in visual outflow velocity implies a pulsation of the outer envelope, on which is superposed a slow outward drift (1965 observations by Doazan, reanalyzed under the stimulus of the modern far-UV observations).

(6) The mass outflow inferred from the visual lines differs from that inferred from the far-UV lines. This implies that the outer cool envelope acts as a valve, regulating the (time-dependent) amount of the mass outflow which becomes mass loss at various epochs. *Mass-loss inferences depend strongly on the region which is observed; one needs simultaneous observations of all regions to be able to map empirically the time-dependent kinematic state; thence infer the equations required to represent, and hence be able to construct a theory for, the dynamic state of the atmosphere as a whole.*

(7) The association between far-UV line components, and V/R of Hα, suggests that the cool

outer envelope is being driven to pulsate—as well as being filled—by the variable mass outflow from the underlying hot regions. Such implied variable momentum input to the cool envelope endorses the variable acceleration of (5.b) above.

(8) Because of the large variability in velocity of these variable components—observed in some stars to range between 200 and 2000 km/s—the simplest acceleration is radiative. Because of the mass-outflow variability, a variability in radiative opacity of the hot regions must occur. Such opacity variability—even with no change in T_e—varies the radiative flux, and therefore varies any acceleration by the hot-region radiation field. No significant variability in the photospheric radiation field has been observed in association with the above-cited variability, of either narrow high-velocity components, or Hα. *Therefore, the presence of a "coronal radiative acceleration" is inferred from these variability studies.* We ask whether it is also generally present in stars not showing such variability.

It therefore appears that some overlap of hot and cold regions along the same radius—with the cold regions occurring farther from the star—is required by the ensemble of observations, especially time-dependent ones, spread over long-time baselines. This suggests, as at least a first approximation, a radial sequence of atmospheric regions similar to the solar, *except* in the presence of a postcoronal decelerated and cooled region. If we include the "dusty regions" demanded by observations of many kinds of stars, including the Bp and some Wolf-Rayet stars; and if we include the solar planetary system as part of the solar atmosphere; we need not even consider the "cool" envelope as exceptional; we simply need more detailed understanding of "cooling mechanisms," just as we need such for "heating mechanisms" in the lower atmosphere. Although, in such a first approximation, the only atmospheric gradient in thermodynamic quantities is radial, the approach only means that the *strongest* gradient is radial—in strong contrast to the historic picture of the cool envelope being confined to being a thin equatorial disk. The pattern and thermodynamics of such a sequence has been proposed by Doazan and Thomas in the B–Be volume of this series (1982); has been elaborated more generally in the volume on *Stellar Atmospheric Structural Patterns* (Thomas, 1983); and is further developed by Cram in comparing models for T Tauri stars to the Sun in the next volume on the F, G, and K stars (1988).

Basically, the modeling scheme follows the solar pattern in the situation of an only mildly varying mass outflow; the Be pattern, for a strongly varying mass outflow whose velocity amplitude is small at the photospheric level; and another type(s?) pattern, whose characteristics are yet to be developed (empirically), for radially pulsating configurations, and for superthermally expanding configurations at the $\tau = 1$ level. (The latter of these links to the configuration of a limiting quasi-thermal photosphere, as already discussed in Section II.) The entire modeling scheme is empirical-theoretical, not speculative-theoretical, in its basis. Hence, it awaits that observational guidance prototyped by the Be studies just abstracted. The Be stars are among the easiest to study because they are bright, their variability is pronounced, and they compose 20% of the B stars. Because the Wolf-Rayet and O stars are less variable, but brighter, they are easy to study. The T Tauri are Be-like in variability, but quite faint, making them difficult to study in that detail necessary to obtain the kind of information just abstracted from Be studies. The Sun, although intrinsically faint, is close enough for detailed study; but its mass-loss variability appears to be small. However, the combined information from all these stars illustrates why we call the empirical/theoretical approach "solar-stellar *symbiotic*."

The above abstract also illustrates why we consider the study of mass-outflow—and mass loss (the two are hardly the same, as above)—*variability* to be the key to not only mass-loss theory, but (conjointly) also to stellar-structural (and evolutionary) understanding and modeling. We turn now to the currently popular mass-loss theory for hot-stars—an abstracted version of the present state of which has been generously (responding to our last-minute plea to bring this volume up to date on its most recent developments) prepared by Kudritzki, Pauldrach, and Puls from their current studies.

III. ONE ASSESSMENT OF THE STATE OF CURRENT HOT-STAR THEORY

A. Empirical-Thermodynamic Demands on Open-System Theory

The first essential thing is to place stellar mass loss into the thermodynamic context of: (a) its observation in hot stars defining them as *open* thermodynamic systems; (2) all well-studied stars—hot, warm, or cold—showing some mass loss. *Therefore, observations demand that any general theory for the structure and evolution of arbitrary stars—not just hot stars and the Sun—give them freedom, and contain mechanisms, to produce nonzero mass flux: to be open systems.* Thermodynamic orientation comes with ability to delineate: which aspects of freedom and mechanisms to produce mass loss are common to all stars; which are peculiar to particular classes of stars; and which arise in individual characteristics of stars. We do not have such orientation from any existing theory: either general theory of open thermodynamic systems; or specific theory of stars as particular types of open systems. We therefore need *observational* delineation of the common, the peculiar, and the individual aspects of mass loss which any adequate theory must represent or predict. The monograph series exhibits our current knowledge of these aspects for stars generally; Chapter 4 of this volume, for O and Wolf-Rayet stars.

Such freedom contrasts strongly to the constraints of classical, speculative-theoretical modeling: under which stars must be closed thermodynamic systems, producing only energy fluxes and forbidding mass fluxes. Although the definition of open versus closed is simply the presence of a mass flux versus its absence, its implication is broader than simply what fluxes emerge from the outermost atmosphere. *Openness characterizes the star as a whole, not just its atmosphere.* Open and closed thermodynamic systems generally differ not only in the existence of a mass flux, but also in that internal plus environmental structure which produces a differentially outflowing, rather than a quiescent, atmosphere, and in that boundary/transition-region structure resulting from it. *We therefore expect differences in our models of internal structure and evolution, not just in outer atmospheric structure, when we acknowledge stars as open systems rather than closed.* Such differences have thus far been only superficially addressed by theoreticians. The effect of mass loss on internal structure and evolution has been confined to only introducing a stellar mass which decreases, empirically, much faster than by nuclear burning.

Again, we do not have insight from existing theory into what differences in stellar structure are to be expected when a star's modeling as a closed system is changed to its modeling as an open system. But here, observations cannot directly guide us; at best, they tell us the star's actual structure, which corresponds to whatever kind of system it actually is—presumably, from observations, an open system. So it is not surprising that comparison of closed-system predictions with observations has produced varieties of anomalies. Any features of agreement between closed-system theory and observations should, in principle—and to the extent of observational imprecision—delineate features common to both open and closed modeling (*not* be considered "gross" confirmation of classical theory). Therefore, both anomaly and agreement provide empirical guidance on those aspects requiring change, and on those which do not, from classical closed-system theory. Excluding peculiar/anomalous, stars/features, from the process of developing, empirically, theory for "normal" stars simply excludes a major part of the observational guidance. We need to synthesize the observed/inferred structure of both normal and peculiar stars into a specification of what any adequate theory must be able to represent or predict. Any "improved" theory of "normal" stars that ignores peculiar stars will be *a priori* incomplete in the preceding sense. The volumes in this series exhibit current observational inferences on actual atmospheric structure across the HR diagram, and its relation to mass outflow; this volume, for O and Wolf-Rayet stars. Opinions differ on atmospheric-structural similarities of solar, hot, and other stars; it remains to relate them all to internal structure and evolution.

While existing theory is inadequate to predict specific structural changes as closed-system modeling enlarges to open, one should expect change to be as basic in stellar structure, evolution, and cosmology as that occurring in the passage from historic effectively-*isolated*-system modeling (e.g., Emden's Gaskugeln) to closed-system modeling. That passage's major aspects were: (1) change from time-independent stars to delineation of only long, evolutionary, time change in a star's structure and appearance, and (2) change from no atmosphere to a single region, directly observable, atmosphere associated with radiative energy transfer.

(1) Emden's focus lay on estimating values of the thermodynamic parameters of stars modeled as the "isolated thermodynamic enclosures" discussed in classical thermodynamic equilibrium studies (e.g., Caratheodory); that is, their masses and temperatures. Both stellar energy production and all stellar fluxes were ignored; they were considered to be second-order effects on such estimates, and stellar evolution therefore fell into the same second-order category. In such first-order cosmology, stars are unchanging objects. But these isolated-system models were useful in showing that stellar interiors were sufficiently hot and dense that thermal-nuclear reactions could occur, producing energy. Therefore, they invalidated themselves if their relaxation times to their extreme-LTE states were comparable to such nuclear evolutionary times. The transition to closed-system modeling introduced second-order cosmology, whose objects change with time, but only quasi-statically, according to the very slow time scales associated with the changes in their mass content by evolution of the *quality* of the matter, as nuclei interact thermally to produce energy.

By contrast, the time scales of possible change in the cosmological objects under open-system modeling can be much shorter for two different, but linked, reasons: (a) because change in their mass content reflects faster mass evolutionary time scales associated to aerodynamic mass outflow, not the slow time scales of mass change via nuclear reactions only; and (b) open-system structure can be nonthermal, and such nonthermal modes can give short-term, nonevolutionary, structural variability. An adequate theory of open-system *structure* must identify: (i) the range of possible nonthermal modes, (ii) which stellar properties determine what modes could be present for a given star, and (iii) hence, what time scales of observable variability should arise from these nonthermal modes, if they exist. An adequate theory of open-system *evolution* must clarify whether evolutionary change in stellar structure comes mainly from: (α) change in the *quantity* of their mass content, (β) change in the *quality* of their mass content, and (γ) change in the nonthermal modes underlying their mass outflow. Therefore, a change from closed- to open-system modeling opens a much wider range of possible time scales for any "change in the cosmological objects" than existed under either isolated- or closed-system modeling. Thus far, hot-star theoretical focus rests on (α)—via luminosity—and (β); (γ) has been ignored because such theory attributes all hot-star mass outflow to the (thermal) photospheric radiation field. Hence, the only variability should be evolutionary, on the long time scales associated with (α) and (β), even if these are shorter than the closed-system variety. The observational contradiction with such a result, from the time scales abstracted in Section II.B, show the present inadequacy of such theory, independent of its details. However, I stress that we do not yet know, empirically, whether the short time scales of the observed changes refer to what is driving the mass outflow or to the atmospheric response to it. Again, the time scales observed in both normal and peculiar stars must be considered in empirical improvements of the theory, hence, in defining that cosmology whose stellar objects are open systems.

(2a) Isolated-system models were even more self-destructive by their exhibiting conditions where radiative, rather than only convective, fluxes must exist at the outer boundary, even to first-order, contrary to their neglect in these models. Isolated-system results therefore demanded *at least* closed-system models. They also differed from a strictly Caratheodory-type enclosure by their radial inhomogeneity; but Emden's imposition of the extreme LTE conditions of polytropic, quasi-static, convective energy-mixing, and the outer boundary conditions of vanishing pressure, density, and temperature, which also imply no interstellar medium (ISM), tried to enforce the "isolated-system" picture. Because such boundary conditions prohibited the model's transition to an ISM, hence the existence of an atmospheric transition region, one could not observationally diagnose atmospheric structure as a test of its thermodynamic consistency. The model's strictures inhibited its wholly theoretical self-checking for consistency. That we verify the existence of stars via their observed radiative fluxes, exhibits the models' inconsistency to whatever level of approximation the modeling of their structure, evolution, and stellar cosmology depends on this radiative flux. The same comment holds on classical closed-system models regarding the kind of description adopted for the thermodynamic state of the atmosphere, and the forbidding of nonradiative energy and mass fluxes. The models' predicted atmospheric structure showed inconsistency with the thermodynamic description imposed, and instability against production of the forbidden fluxes; whereas the size of the observed mass fluxes imply faster mass evolution than the models' evolutionary description imposes.

(2b) In essence, the motivation for closed-system modeling of stellar structure arose in the demonstration (Schwarzschild-Eddington) of the greater importance of radiative, over convective, energy transfer throughout most of the star, but especially regarding energy escape near the outer boundary. The isolated-system models were unstable against radiative energy fluxes. Modelers tried to correct this instability by admitting a radiative energy transfer throughout the star, and a radiative flux at the boundary; but permitting no other fluxes: thus speculatively defining the star as a closed system. They continued the isolated-system thermodynamics of modeling internal energy states, and defining the relevant thermodynamic-state parameters, via speculatively imposing LTE.

Thus, the greatest change in stellar structural models occurred in the outermost regions: a radiation-conditioned atmosphere arose as the transition to an optically transparent, local, ISM which was fed by a radiative energy, but not mass, flux from the star. An LTE description of the ISM's thermodynamic state was not forced on it; there was no discussion of transition from the star's LTE to the ISM's non-LTE. The "transition parameter" of the atmosphere was the opacity, τ; the stellar "boundary" signified the vanishing of τ, not of T; and a nonvanishing "LTE boundary temperature" provided thermal radiation; the contribution of each atmospheric element being the LTE emissivity times the element's opacity. This thermal radiation balances the thermal nuclear-energy production in the interior; thus remedying, in principle, the basic inadequacy of isolated-system modeling. Thermodynamically, this radiative-transfer-dominated atmosphere is a single region in radiative and thermal-hydrostatic equilibria; the distribution within which of its LTE-state variables is fixed only by the three parameters: effective temperature (radiative flux), gravity, and chemical composition (mass quality). At least approximately, its two-dimensional (T_{eff},g)-dependence corresponded to the observed "main-sequence" of stars on the observed HR diagram, and "closed-system evolution" appeared to represent the giant/supergiant branches. Again, the basic questions were: (a) the significance of peculiar stars, and (b) the thermodynamic self-consistency of this approximation (now the closed-system) even for normal stars, especially in their outer regions. Is isolated—but inhomogeneous—system LTE valid near optically thin, low matter-density boundaries of closed-system stars? We also ask: do such boundary regions damp or amplify changes in atmospheric structure by other fluxes than radiative? That is, does closed-system modeling also lack self-consistency when applied to stars?

Fifty years after closed-system modeling was introduced, we demonstrated—empirically-theoretically, using solar observations + a thermodynamics free of imposed LTE—that an LTE treatment of the outer atmospheric regions is thermodynamically self-inconsistent and badly overestimates the radiative fluxes for given $T_e(\tau)$. This makes difficult the diagnostic testing of the assumption of RE, even in wholly thermal atmospheres, until LTE is replaced by fully non-LTE models. Thirty years later, RE has still been tested in satisfactory detail only for the quasi-thermal Sun, where RE is found to be invalid for $\tau_C < 10^{-3}$, thus to be invalid in regions where the centers of most strong lines are formed.

Shortly thereafter, Parker (1958) showed that the nonradiatively heated solar-coronal regions were unstable against aerodynamic mass outflow, resulting in mass loss, hence defining the Sun as an open system. Much later, Cannon and Thomas (1977) showed that *any* HE + RE atmosphere is unstable against amplifying any small, outward, radial velocity; that the outflow is unstable against producing acoustic waves when it reaches almost thermal velocity; and that a monotonic outward rise in T_e accelerates the mass outflow to a mass loss. It remains a debated question whether the nonradiative energy flux, and its dissipation, are causally linked to, or only associated in origin with, the mass flux. (I note again that the maximum outflow velocity, but not the size, of the solar mass outflow varies by a factor of 2 in evolutionary short time scales. For Be stars, both vary, and strongly.) In any event, both nonradiative fluxes arise, and the only association with the *visual-spectral* radiative flux discovered thus far is a *decrease* in visual + far-UV luminosity as mass outflow increases, for one Be star, at one epoch. (More observations, simultaneous in the far-UV + visual, are badly needed.)

Thus, other atmospheric regions than the closed-system photosphere exist, and are the locales where originate diagnostically-important, observed, spectral features. Empirically, one could conclude only that closed-, like isolated-, system modeling fails to predict or represent all the thermodynamic fluxes which arise in actual stars; and that, in consequence, it fails to predict or represent those exo-photospheric regions associated with these fluxes. One can dispute whether the structural details of Sun and other stars are similar; but not whether the basic thermodynamics must be uniquely different for the Sun than for other stars, in order to "uniquely" interpret or explain features which its unique location lets us observe in it, *but not yet* in all other stars. Any improved/enlarged, broadly applicable, thermodynamically-sound theory must be able to describe *all* fluxes from *all* stars and their relation to stellar structure and evolution, or likewise be judged only incomplete.

Consequently, for the Sun and any other stars, hot or cold, we faced a situation in which even neoclassical (i.e., including non-LTE) standard closed-system modeling failed to represent or predict: (a) the observed existence of non-RE atmospheric regions; (b) the inferred existence of energy fluxes other than radiative; (c) the consequences of such atmospheres being unstable against a small-velocity mass outflow somehow arising in the deep atmosphere, to produce a mass flux; and (d) how/why such outflow arises in the deep atmosphere. Sections I and II of this Perspective abstract the evidence for similar phenomena occurring in hot stars. Any theory, therefore, which accepts such stars as open systems gives only an incomplete representation of such stars unless it includes *all* the summarized observational phenomena as features which the theory must incorporate. Again I stress: *features of peculiar, as well as normal, stars.*

To put proposed theories of hot-star mass outflow—particularly its origin—into focus, it is useful to elaborate this normal/peculiar star comparison. The basic question is whether they are thermodynamically-different kinds of objects, or thermodynamically-similar objects that differ only in flux "amplitudes" and in flux time-dependence. We ask what is the range of open-system "individuality-diversity."

Then the focus of *isolated*-system modeling is on the mass, and thermal energy, content of stars; fluxes are ignored; stars do not change. The focus of *closed*-system modeling is on the relation between the mass content of stars and their luminosity (i.e., radiative flux for given radius); stars change, but the time scale is very slow, with evolution in mass *quality*. The focus of *open*-system modeling is on understanding, symbiotically, the relation between size, origin, and time-dependence of all three fluxes (radiative energy, nonradiative energy, and mass), and the stellar structure (interior and atmospheric) produced by a given quantity and quality of matter—with its individual macroscopic (angular momentum, etc.), as well as microscopic (nonEquilibrium populations of matter and energy states) properties—which condensed to form the star. To first approximation, open-system time scales are evolutionary, simply shorter than closed-system, because aerodynamic mass loss replaces nuclear-burning mass loss. To higher approximation, admitting a possible diversity of nonthermal modes, for similar values of closed-system parameters, open systems can produce the observed range of even shorter time scales, which *seem* to contradict even open-system evolutionary change.

This contradiction implies nonthermal-mode structure for stars; the question is their amplitudes. Need they be the same for all stars having similar values of the closed-system parameters? Given the instabilities inherent in even quasi-thermal models, any improved theory should focus on, or at least include, such nonthermal modes, interior as well as atmospheric, and on the question of the possible range in their amplitudes for given values of closed-system parameters. *Modeling stars with similar values of closed-system parameters, but with different amplitudes for those nonthermal modes permitted in open-system stars, is basic to understanding the normal/peculiar star difference.*

The basic assumption that closed-system modeling can produce stars which are time-independent, except on nuclear-evolutionary time scales, rests on the ability of stars to "find" a stable configuration whose internal energy production is balanced by the radiative-energy flux from the atmosphere. Failure to find such a stable configuration must result in either or both of the following: (a) other kinds of energy fluxes from the atmosphere, requiring nonthermal mass motions; (b) changing the energy production in the interior which, given the similarity in central conditions for all quasi-thermal modeling, again requires a nonthermal structure (now for the interior) with consequent nonthermal atmospheric effects.

That most stars exhibited, from visual-spectral observations alone, neither (a) nor (b) gave some astronomers confidence in the concept of "normal" stars, with structure and evolution following the closed-system pattern. That an increasing number of stars were peculiar or anomalous in the sense of the preceding summary stressed the point made above: the similarities between closed-system models and real stars show the common features of closed- and open-system modeling; peculiarities delineate the differences. For astronomers who adopt this viewpoint (mainly observationalists), a focus on only "normal" stars was hardly adequate in researching either the *facts* of real-star structure or the thermodynamic basis for its general modeling. Hence, one finds the extensive literature—over the last century—on the observational study of peculiar stars. Unfortunately, it has not been paralleled by corresponding theoretical activity on expanding normal-star modeling to include peculiar stars.

Contemporary observations outside the visual spectrum substantiated that outlook by showing even "normal" stars to violate closed-system strictures; for these stars, violations become detectable only in those exophotospheric atmospheric regions observed in the far-UV. Nonthermal amplitudes increase toward the (free) boundary—conditioned by the amount of dissipative damping. So "violation" is a matter of degree; we must ask how and why some stars show stronger anomaly/peculiarity than others, detailing the aspects. Relative to some current thinking on hot stars, we particularly ask if and why some stars depart from closed-system models only by the presence of a mass flux, somehow obviating the general open-system presence of nonradiative energy fluxes and dissipation. Or is such modeling only incomplete theoretical, which does not ask into their models' instability against nonradiative energy dissipation?

Prior to far-UV spatial observations, some astronomers hoped that some (single) "mildly peculiar" stars, showing very short time changes but not "cataclysmic" ones (i.e., not mass ejection), would be eventually incorporated into closed-system modeling. Eddington's exploration of closed-system internal structure, and radiative energy transfer, originated in such observed nonthermal stars—the pulsating variables. Closed-system modeling was apparently able to embrace such stars by recognizing "opacity-valves," associated with radiative energy transfer, which: permit small density perturbations to amplify, change the opacity, change the amount of energy stored locally, and introduce phase lags between energy production and radiation. Under linear theory (ignoring pulsation amplitude), this could be done without violating closed-system boundary conditions. The atmospheres and subatmospheres were nonthermal, but *apparently* could be modeled as closed systems, with only radiative energy fluxes. The stars were indeed only "mildly peculiar."

However, when one asks what fixes the pulsation amplitudes (i.e., limits them from growing so large that the star becomes unstable), one recognizes that the large velocity gradients accompanying velocities amplified by the atmospheric density gradients produce shock waves and energy dissipation. Current thinking therefore finds nonradiative energy dissipation to be the amplitude-limiting process, producing the pulsating-star equivalent of the solar chromosphere, but something which has not the quasi-static, thermal HE, character of that solar region.

In addition to this scenario applying to a classical, cool-type, pulsating star, it has also been suggested to be the actual configuration for a very massive hot star, which historical modeling presumed could not exist because radiation pressure prohibits its formation. However, the same current thinking suggests that the radiation pressure would simply decrease the effective gravity enough to make the star unstable against pulsation but, as above, with amplitude limited by nonradiative dissipation and possibly mass loss. Therefore, we should admit the possible existence of such stars—more massive and hence hotter than nondissipative closed-system modeling would predict—as open-systems. However, their mass loss would not be driven by radiative pressure/acceleration but by pulsation (cf. suggestions by Davidson and Humphries (1982); and by Willson and Bowen (1985)).

Finally, one notes that current thinking seeks the source for the nonradiative heating of at least the low solar chromosphere in dissipation of those waves produced by either solar nonradial pulsations or time-dependent convection, the diagnostic distinction between the two being difficult. The question of how far into the chromosphere/corona such nonhydromagnetic energy transfer and heating suffices to maintain that energy balance which produces the empirically inferred $T_e(r)$, apparently depends on resolving the nonlinear coupling among the modes of the wave spectrum—their nondissipative and dissipative outward propagation (Hammer, 1985). Thermodynamically, therefore, the similarity between the variety of pulsating stars and the Sun, regarding the origin of the nonradiative fluxes which give all these kinds of "peculiar" stars their open-system character, is strong and basic. It is the existence of any variety of subatmospheric pulsational (nonthermal) modes—whose amplitudes are fixed by nonlinear mode-coupling, propagation, and dissipation in the atmospheric boundary/transition regions. One just begins to follow these directions of study in hot stars (cf, the reviews by Baade and Henrichs in this volume).

So, as in the LTE/non-LTE problems in the boundary regions of the atmosphere, one asks stability against nonlinear effects: there, in thermal radiative transfer; here, in nonthermal mass transfer. As abstracted previously, the original "explosion" in variety of regions comprising the solar atmosphere; and the recognition that this boundary/transition zone is unstable against mass outflow; were each associated with an observed nonradiative heating whose origin was sought in various aspects of nonlinear closed-system instability against the production of nonradiative energy and mass fluxes. It would seem that now, in developing new open-system theory for mass outflow and the associated atmospheric structural patterns, a major theoretical focus should lie on the instabilities of any open-system modeling that includes *only* a mass outflow/flux, especially a time-independent one. If such instabilities exist, producing either nonradiative heating or a time-dependence, the resulting self-consistent atmospheric structure will differ significantly from that of current theory.

B. Comment on Current Hot-Star Theory from the Standpoint of Satisfying the Preceding Empirical Demands

Based on the viewpoint of the discussion in Section III.A, we divide current explorations of mass-loss modeling and the associated atmospheric structure for hot stars into two broad approaches: (a) those which attempt to predict the value of the mass flux and associated atmospheric structure from values of the other fluxes, gravity, and composition—thus far via time-independent modeling; and (b) those which use empirical values of the mass flux, and other fluxes + gravity, ignoring composition, to model atmospheric structure, admitting time-dependence. Here, we focus on (a); (b) is illustrated in the previously abstracted Be modeling by Doazan and colleagues.

1. Theories Focused on Predicting Mass Fluxes from Other Quantities. This approach corresponds to simply freeing closed-system, time-independent, spherically symmetric, normal-star modeling from its outer boundary condition of zero mass-outflow velocity—by giving the integration constant of the equation for conservation of mass some nonzero value. It remains only to specify the acceleration and any nonradiative dissipative processes admitted throughout the flow regime/atmosphere. The closed-system radiative-transfer equations, and the rate equations describing the populating of microscopic energy levels, are modified to include the effect of moving rather than static medium. This approach is best characterized as computer-*stimulated* modeling; because once the acceleration, nonradiative dissipation, and integration constant of the mass-conservation equation (i.e., the rate of mass loss) are specified, it is an initial-value problem—from prescribed values of radiative energy, nonradiative energy, and mass fluxes in the deep atmosphere—to compute those radial distributions of velocity, thermodynamic state-parameters, and microscopic energy-level populations which we cannot yet directly observe.

The computing problem is hardly routine; it involves the coupled mass and radiative transfer, and microscopic rate, equation. More difficult, thermodynamically, however, are the problems of choosing the dissipative terms, the accelerations to include, and the value of the mass loss. Astronomers have habitually bypassed the first problem by parameterizing its result—prescribing a $T_e(r)$. The custom is to choose the acceleration as a function of star type, based on some hypothesis as to what it is, rather than attempt to infer it from the observed velocity field. (One must admit that, before space observations in the far-UV, there was little observational guidance; but now the situation has changed.) From these two assumptions on two of the basic three quantities entering the adopted equations, the custom is to try to predict the third—the mass outflow/flux. Because one adopts vanishing small outflow velocity in the lowest atmosphere, density and T_e distributions there are unaffected by the mass outflow. One can obtain them from closed-system modeling *if*, as is customary, one assumes the absence of nonthermal modes in parallel to assumed predictability of the mass loss via the stated procedure. Because, under these conditions, the value of mass loss and that of *initial* velocity in the deep atmosphere are equivalent quantities, one must be careful in beginning the outward tracing of the flow pattern.

My own viewpoint, which I have elaborated at length elsewhere (Thomas, 1973, 1983), is that the mass outflow in the low atmosphere must be taken as an independent, time-dependent parameter whose value may ultimately be predictable from that complete theory of stellar structure which includes nonthermal modes. But I do not discuss the basis for this approach here, only to note it for contrast.

It is useful to place the original solar-wind theory into this context because its approach was also that of predicting the mass loss by an algorithm *imposed—not proved* to be valid—on the outflow configuration in a way resembling that adopted in popular hot-star theory. This solar theory was initially, *a priori*, rejected for hot-star application, by the originators of the hot-star wind theory (Lucy and Solomon, 1970) as requiring too large T_e values (some 10^7 K) in hot-star atmospheres. Modern X-ray observations require such T_e *if* the X rays are thermal. Current hot-star theory considers the main body of the outflow to be unheated; the X-ray emission coming from isolated, but quasi-uniformly distributed, and small, "hot pockets" in the flow, hypothesized as resulting from shocks produced by local instabilities in the "hyperthermal" outflow. Such instabilities, in monotonically accelerated superthermic outflow, have not yet been observed in the laboratory.

As we have earlier mentioned, dissipative instabilities in steady flows accelerated from low subthermic velocities to hyperthermic are, in the laboratory, more usually observed to occur in the transthermal region. We exhibited (Cannon and Thomas, 1977) their stellar atmospheric counterpart, in asking the stability of flows accelerated only by the density gradient and gravity and heated both radiatively and nonradiatively. These considerations have been largely ignored by the hot-star theoreticians. But at least, because of the X-ray evidence for *some* nonradiative heating, producing *some* regions of $T_e >> T_e$(RE) in *some* hot stars, current theories of hot-star structure are becoming less *a priori* restricted, and more influenced by observations. The currently-most-popular focus on modeling only those configurations in which there is no nonradiative dissipation—arising either independently or from instability of the mass outflow—is now evolving to include *some* types of dissipative instabilities, as above. However, to preserve the basic theoretical picture of a mainly nondissipative mass outflow, theories currently explore only local, rather than spherically symmetric, dissipative instabilities in *only* the superthermic regions of an otherwise spherical, time-constant, quasi-RE mass outflow. In a wide variety of other stellar types, especially in the Sun, nonradial atmospheric inhomogeneities have been historically explored to represent observations which appear to be anomalous under whatever strictly spherically symmetric models are in use. However, as solar-illustrated below, today's modeling of outflowing atmospheres cannot ignore the "perturbation" introduced by the distribution, and structure (time, space), of the transthermal regions in discussing the "predictability" of the size of the mass outflow. We, therefore expect hot-star modeling to evolve further, toward greater transthermal detail.

Such shock-wave heating in the superthermic region as that sought in the above local instabilities is indeed to be expected in a *variable* mass outflow, and its resulting nonmonotonic radial distribution of velocities (e.g., Be modeling, abstracted previously). However, this does not produce the "hot-pocket" type of heating; at most, a pole-equator asymmetry results from including a rotation effect on the deceleration arising from the mass-outflow variability. Such variable mass outflow is integral in the empirical approach (b), not in the present wholly theoretical approaches (a), which as yet include no mechanisms to introduce the variety of observed variability abstracted previously.

So, to put into perspective that hot-star approach to "predicting" mass loss, which explicitly excludes the presence/effect of any nonradiative energy fluxes, we must be very clear on that solar algorithmic approach to "predicting" the size of a mass outflow which explicitly depends on the existence of such nonradiative energy fluxes. Comparison of the similarities and differences of the two approaches makes very clear the basis for thinking one can predict the mass flux from some combination of the other fluxes and the gross stellar parameters such as mass and gravity—without evoking a nonthermal structure for the star: subatmospheric and atmospheric. It also becomes clear why modern observations negate this basis. That is, we see why both closed-system, and thermal open-system, stellar modeling—in their requiring a thermal structure for the atmosphere and subatmosphere—are inadequate to represent the range of modern observations.

a. Perspective on Mass-Loss Prediction via the Solar-Wind Algorithm. The original solar-wind theory (Parker, 1958) was constructed from empirical T_e and density distributions in the corona and beyond. These showed that the corresponding one-dimensional thermal velocity, q, is large enough, at radii where significant matter concentration still exists, to permit that matter to escape from the gravitational field—and form that solar particle flux long inferred from its terrestrial and cometary effects. Early discussions aimed at clarifying whether such mass loss was by quasi-static thermal evaporation from, or thermal aerodynamic expansion of, the super-hot corona. To escape, an evaporating noninteracting particle's velocity must exceed the local escape value. However, a gravitationally bound atmosphere has a pressure gradient, which provides an acceleration that balances gravity under HE, and exceeds it under those conditions supporting a mass outflow. So, an aerodynamic outflow can, at given r, have a subescape velocity, yet result in a mass loss, utilizing the pressure gradient for the needed acceleration. The essential questions are what pressure gradient arises in a given star, and what size mass outflow can it accelerate to escape as a mass flux from the star. An answer requires a complete expanding-atmosphere model, which is aerodynamically self-consistent between the size of the mass outflow and the velocity distribution in all the regions from the subatmosphere out into the local environment. If the gas-pressure acceleration does not suffice, and a mass flux is observed, one seeks other alternatives.

The significant aspect of Parker's algorithm is that its construction rests on considerations of the coronal and postcoronal regions only. Of the lower regions, one demands only that the outflow velocity become negligibly small with depth. According to the strictures of his algorithm, modeling begins at the thermal point, not in the subatmosphere. I abstract this approach; it puts into sharp focus the suppressed role of the transthermal regions in their relation to the "prediction" of $\dot{M}$ and $v(r)$.

Parker's algorithm then rests on applying the two conditions: (1) for an expanding outflow to be accelerated rather than decelerated, its velocity must exceed the (one-dimensional) thermal value, q; (2) under the (solar-type) condition that the acceleration be due to gravity and gas pressure, such accelerated expansion becomes a mass loss if it passes from subthermal to superthermal velocity at that radius where q (which therefore equals the outflow velocity) equals half the (local) escape velocity (so the total *mean* one-dimensional particle velocity equals the escape velocity). The algorithm for "predicting" the size of the mass loss is as follows:

(a) From the empirical $T_e(r)$, find the lowest r satisfying $q^2(r) = GM/2r$.

(b) To evaluate the mass loss, $\dot{M} = 4\pi r^2 \rho q$, at this r and q, adjoin an empirical $\rho(r)$.

(c) The velocity distribution comes from integrating the flow equations under the conditions that the flow be everywhere subthermal below the escape point, and everywhere superthermal above it.

The dependence of the algorithm on nonradiative heating is clear: the distribution $T_e(r)$ fixes the escape point, and is the parameter of the flow solution. That the value of $\dot{M}$ comes from "empirical evaluation" rather than "prediction" is equally clear. One has still insufficient knowledge of the source of the nonradiative heating to predict $T_e(r)$. To predict $\rho(r)$, one requires a solution of that aerodynamic mass-outflow problem whose parameters are the mass loss, $\dot{M}$ (or an initial velocity and density at some r_0), and the distribution, $T_e(r)$, under physically self-consistent, not arbitrarily imposed, boundary conditions. The algorithm for the size of $\dot{M}$ simply applies the condition (2) via (b), using the *empirical* data. Whether the derived $v(r)$ is correct rests on the validity of the boundary conditions (c), and this depends on their self-consistency with the value of $\dot{M}$

It is clear that the solution, $v(r)$ and $\rho(r)$ for prescribed $\dot{M}$ and $T_e(r)$, must satisfy the condition $v(r_c) = q(r_c)$ at the "critical/escape point r_c" (condition (1) above). It is, however, not clear that $v(r_i) = q(r_i)$ does not occur several times, each at a different radius $r_i < r_c$. The only requirement for such a possibility to occur is that, in each case—except the last, $r = r_c$ — $v(r_i + \delta r) < q(r_i + \delta r)$. This condition is satisfied by admitting a shock wave between r_i and $r_i + \delta r$. The shock both decelerates and heats. *A condition imposing that outflow velocity reach q for the first time at the escape point restricts the solution to shock-free flows* (i.e., to the configuration where r_c is the lowest-lying r_i). Thus, for a given $T_e(r)$, the value of $\dot{M}$ must be such as to produce this location as the "first" thermal point, *if* the condition (c) is to be valid.

We have already noted that—because $\rho(r)$ is HE for $v << q$—this condition on the value of $\dot{M}$ is one on v in the low HE photosphere and on $T_e(r)$. Clearly, such a condition gives only a minimum value for $\dot{M}$: that which, from an HE solution under the prescribed $T_e(r)$, produces $v \sim q/3$ at approximately one density scale height below the escape point. Any larger value of $\dot{M}$ produces a $v(r)$ in the HE region that can also satisfy condition (2) of the algorithm, *but* it requires the above sequence of shocks. Hence, it can be compatible with algorithm conditions (a) and (b), but not with (c). By construction, Parker's algorithm produces the actual value of $\dot{M}$, but not necessarily the correct $v(r)$. The correct transthermal regions are not reproduced for an $\dot{M} > \dot{M}$(min). Actually, the entire sequence from about one density scale height below the first shock through the critical point is the "extended" transthermal region. This means that the $v(r)$ predicted under the imposed shock-free solution will be everywhere too small, hence $\rho(r)$ everywhere too large, below the critical point.

The driving source of the outflow is clear: the local thermal energy. In the lowest, RE part of the atmosphere, the thermal energy is fixed by the star's radiation field; in the non-RE regions, it is fixed by the star's radiation field + the nonradiative dissipation. Therefore, given a somehow imposed, small, outward initial velocity, the instability amplification mechanism is indeed a "radiatively-driven" acceleration; but arising energetically, not by radiative-momentum, transfer. The latter, with Doppler-shifted desaturation of the line-driving, is assumed to be the acceleration in hot stars. But again, one requires some source for an initial velocity, and the gas-pressure gradient for the initial acceleration. Without such an initial velocity, the atmosphere remains in the (unstable) closed-system, thermal, HE configuration. We ask if it suffices, to be open, that all gravitationally bound stellar atmospheres are unstable against *any* radial-velocity perturbation—depending on the dissipative mechanisms, including mass loss, to fix instability amplitudes anywhere?

Thus, the original solar-coronal theory of mass outflow restricts a priori the flow to being "perfect" (i.e., shock and dissipation free). In the laboratory, one designs the throats of supersonic wind tunnels to provide such flow. With such design, for quasi-static reservoir conditions, the mass outflow is fixed by the reservoir and environmental pressures. One can say the mass loss is "pulled out" by the low environmental pressure, as one can say a cold exterior "pulls out" heat from a reservoir. But unless the wind tunnel is "driven" by refilling its reservoir, it is transient: the pressure excess of the reservoir, and outflow from it, decrease with time; the nozzle is no longer "perfect." A driven (continuous refilling and outflow) tunnel has a velocity imposed at the base; it is "partially pushed from below"; its reservoir velocity and mass outflow are specified. If the same geometry is used for the nozzle as for the static case, thermal velocity is reached earlier, and a shock arises. The outflow is dissipative, and "driven out" by the nonthermal state of the reservoir. To obtain a shock-free flow, one redesigns the nozzle geometry. *A priori, there is no reason to expect the star to design its subatmosphere/atmosphere/local environment structural complex to produce a shock-free perfect-nozzle outflow*; especially when so much evidence for dissipative flow exists. Therefore, one must be cautious in applying any theory based on such restriction, to "predict" the mass loss from a star modeled as having a thermal subatmospheric + lower atmospheric structure.

Current solar modeling focuses on local hydromagnetic instabilities and heating above the middle chromosphere; and the distinction between coronal "holes" and non-holes as magnetically open and closed regions. The situation is hardly clear at the moment (cf. Jordan, 1981; Lites, 1985). So the preceding remarks are solar-applicable only to the extent that the original solar-coronal modeling of mass loss is still of some interest. Their aerodynamic cautions are however useful in examining the self-consistency of current hot-star attempts to *predict* mass loss.

For aerodynamic discussion of such multishock nozzles, see Liepmann and Puckett (1947). For laboratory examples of "free discharges" producing the above periodic compression/expansion pattern, see Prandtl (1904) and the numerous U.S. Army Ballistic Research Laboratory (Aberdeen) photographs of "flying wind tunnels" of various designs (Braun, Charters, and Thomas, 1945). For an application of Prandtl's work, with gravity added, to solar "jets," see Thomas (1950).

b. The Hot-Star Configuration. Then, hopefully thoughtful from the solar example, we turn to current hot-star theory as it has been developed, via *a priori* strictures, in an attempt to predict the size of mass outflow, the corresponding velocity field, and the coupled atmospheric structure, superposed on a thermal subatmospheric structure. As in solar modeling, the problem of nonradiative heating is bypassed by arbitrary choice of $T_e(r)$. Unlike the solar choice—which tries to represent the existence of that large nonradiative heating which produces a steady outward rise in T_e to coronal values—hot-star modeling tries to represent a near-RE situation for the first approximation $T_e(r)$. Also unlike early solar modeling, $\rho(r)$ comes from self-consistent solution with $v(r)$. To the accelerations of gravity and gas pressure, the theory adds that of the photospheric radiation field as perturbed by a mass outflow which is optically thick in some (far-blue) spectral regions. But, to first approximation, the theory assumes no strong increase in the radiation field due to nonradiative heating: the star's "luminosity" is the photospheric one. Because subatmosphere and low atmosphere are assumed to be thermally-structured, the values of mass-outflow size and velocity would be equivalent in specifying initial conditions *if* these were taken to be in the low subthermic regions of the outflow, which they are not—thus, in this, resembling classical solar theory.

The current hot-star theory was originated by Lucy and Solomon (1970); greatly expanded and detailed by Castor, Abbott, and Klein (1975); even more expanded by Abbott (cf. his most current review, 1985); and further enlarged, to exhibit interesting new aspects—especially of the outflow's great far-UV and XUV opacity, which greatly increases the extent of the diagnosable atmosphere (given observations in these spectral regions)—by Kudritzki, Pauldrach, and Puls (cf. their review in Chapter 4 of this volume). The unifying character of all these versions of the theory is their omission of any terms producing any nonradiative heating, and their steady-state character. These two aspects are linked, especially in the regions of transthermal flow. It is, however, not obvious that the time scales of that variability discussed in Section II.B do not permit treatment under a mass outflow which varies much more slowly than the relaxation time for a steady-state flow to adjust to variable lower boundary conditions. In this case, any time-dependence would be only local: in the transthermal region, and in the decelerating and pulsating regions of the cool Hα envelopes, etc. None of the cited versions of the hot-star theory address these observed "anomalies" of Section II.B.

So in the following, just as in the preceding critique of the classical solar approach to mass loss, major focus lies on making explicit just what the existing "hot-star algorithm" for "predicting" mass loss and $v(r)$ actually predicts and what it restricts, on the breadth of stellar atmospheric structure. The summary by Kudritzki, Pauldrach, and Puls in Chapter 4 of this volume gives a good statement of the outlook of the active proponents of existing hot-star theory. So, I comment from the basis of their picture of existing, nondissipative, steady-state theory of *photospheric* radiative acceleration/origin of mass outflow.

In the sense of this theory, "photospheric" includes that radiation produced in regions much above the visual-spectral photosphere, in the radiatively opaque wind, but not necessarily from nonradiative heating. Such radiation must be, conceptually, distinguished from that arising from a chromosphere/corona, whose non-RE origin is unambiguous, but which may also be opaque. The uncertainty is whether any "mild" departures from RE, which may or may not be required by the present results of Kudritzki's group, forecast stronger T_e rises in outer regions, as in the solar case. If so, it is not only the increase in photospheric "hotter *quality*" radiation from the T_e rise that concerns us in a "broader" radiative acceleration theory, but also that change in "hotter *quantity*" of the radiation field that arises simply from opacity changes. Such opacity changes accompany T_e gradients, as well as those density gradients arising from the aerodynamic outflow. They are particularly important in considering the effects of variability—whether the variability is in mass outflow or in nonradiative heating—on those atmospheric features which seem most readily modeled as associated with both variability and with some kind of radiative acceleration.

At the moment, one of the strongest a priori reservations in applying "standard" radiative acceleration theory to general stellar modeling lies in the lack of any observed variability in photospheric radiation at times when variability is observed in some atmospheric structural features which should be linked to radiative acceleration. However, it is not obvious that radiative acceleration is restricted to that arising from RE regions, if non-RE regions exist. We note that the original proposals of a solar radiative acceleration focused specifically on those spectral regions now recognized as chromospheric in origin. Our proposals for the variably accelerated mass outflow driving the Hα envelopes of Be stars invoke their coronal radiation field. Eventually, therefore, we must also ask what changes in current theory are required if it is to be expanded beyond photospheric radiation fields.

In Chapter 4, Kudritzki et al. try to put the essential physics of current hot-star theory into perspective by discussing, initially, a supersimplified version of the theory as highlighting its important features. I follow their approach to equally highlight what the radiative acceleration "algorithm" actually rests on, and "predicts" regarding the mass loss from the star, in a way similar to the preceding critique of the hot-corona theory algorithms. The approach puts into focus what is needed to expand *photospheric* radiative acceleration to embrace variability. Any coronal radiative acceleration is ignored here. For coherence, the discussion follows the section and equation numbering of Kudritzki et al. in Chapter 4.

b.1 Supersimplified (SS) Theory. The authors tell me that their objective in this presentation is to be only heuristic, not to be rigorous. They try to "picture" to the reader why one should expect: (a) dM/dt to be fixed (mainly) by L, rather than its being an independent thermodynamic quantity parallel to L; and (b) the velocity (accelerated outward) distribution to be fixed only by gravity. I appreciate their objective; too often, giving *only* a computing code brings little physical understanding. In the present case, however, their exposition is circular, with the "derived" relation between L and dM/dt being only a tautology; disguised by introducing a parameter, ϵ, which is unjustifiedly taken to be constant among the stars considered. And a gravity-*like* velocity distribution is imposed by the assumed form of the particular accelerations introduced into these approximations to the governing equations. The "derived" results on this imposed r-dependence concern only its coefficient, which—the authors erroneously assert via this ϵ and an inaccurate abstract of observations—depends, to good approximation, only on the star's (photospheric) escape velocity. This is factually incorrect, as we show below.

Thus, their attempts to demonstrate the validity of (a) and (b) above from this SS approximation are misleading, *especially heuristically*. However, their championing of the "predictability" thesis via this SS example, and our following pragmatic critique of its illogic, put into very clear focus: (1) the similarity of solar and hot-star theory in confusing prediction of lower atmosphere mass outflow with computing how much of it can be accelerated to escape, (2) the assumptions necessary to convert a homomorphic association between L and dM/dt into an isomorphic one, and (3) hence, the concepts implicit in seeking to predict the observed mass flux from the observed radiative flux *via considerations of the atmosphere alone*, instead of treating the two fluxes in parallel, each produced by the thermodynamic configuration of the star's structure as a whole. We introduce this focus: (α) in terms of the physical picture implied by the terms retained in the equations defining the SS approximation, and (β) in terms of the solutions of these equations for $v(r)$ over the range of "permitted" $\dot{M}$.

α. *Physical Picture*. Then, as the authors remark, *both their SS Equation (2), and their "less simplified" (S) Equation (21), treat the mass outflow as an ensemble of noninteracting particles, moving under gravitational and radiative accelerations only, within the atmospheric regions to whose description these SS and S equations are applied.* They define the physical characteristics of these regions to be those under which the terms retained in the SS and S approximations dominate in determining the solution of the authors' form of the general aerodynamic equations containing a radiative acceleration. That is, throughout these regions it is imposed that: (i) the outflow is sufficiently superthermic that $q^2/v^2 << 1$, (ii) the thermal velocity is small enough compared with escape velocity that $q^2 << 2GM/r_x$, where r_x is the outer boundary of these regions.

Such approximation excludes "cooperative" thermodynamic effects of gas-pressure gradients, any kind of aerodynamic energy dissipation, any local flow deceleration and heating by shock waves, and thus, any coupling between macroscopic flow and thermal energies of the gas. It therefore excludes those effects by which: (α) solar mass-loss theory was "converted" from thermal evaporative to aerodynamic in the outermost atmospheric layers, and (β) the mass contents of outer and inner atmospheric regions are coupled by aerodynamic mass outflow rather than by noninteracting particle orbits. But in adding the forces converting thermal evaporative orbits to radiative expulsive, this approximation to hot-star mass loss also restricts the radiative acceleration to act individually on particles. "Easier accelerated" atoms are forbidden (collisionally) to "drag along" the less favored—so we must admit the possibility of an inhomogeneous wind structure induced by "spectroscopic configuration sorting." Therefore, in terms of coupling between micro- and macro-scopic aspects, this variety of "effectively rarified" gas dynamics more resembles the collision-free plasmas governed by magnetic fields than it does the relatively high-density (compared to solar) hot-star winds we observe.

That the particle orbits—hence mass outflow $v(r)$—follow a gravity-like, r^{-2}, acceleration within the region is hardly surprising in terms of the imposed characteristics of the region. The only two accelerations admitted are gravity, varying as r^{-2}, and radiation, which the authors represent, by their Equation (1), as:

$$g_{\rm rad} = (K/\dot{M})\ v dv/dr = \text{constant} \times \text{inertial acceleration} \qquad \text{(a) or KPP[1]}$$

where

$$K = L\ N_{\rm eff}/c^2 \text{ and } \dot{M} = dM/dt \quad .$$

$N_{\rm eff}$ is a parameter representing the "effective number of accelerating spectral lines." (Physically, $N_{\rm eff}$ should be peculiar to a given ion in a given region of a given atmosphere, thus giving the possibility of differential acceleration already mentioned. However, we follow their procedure, which adopts one $L/N_{\rm eff}$ value throughout a given atmosphere.) Hence, the only terms *permitted* to enter the SS equations are proportional to either r^{-2}, like gravity, or to vdv/dr, like the inertial acceleration; so a gravity-like velocity distribution $v(r)$ *is a priori prescribed*—independently of the value of L, $\dot{M}$, or whether or not a relation exists between them. (By adopting the usual representation of the effect of a continuum radiation pressure as [the constant] Γ times gravity, they impose an r^{-2} dependence on it.) Therefore, their adopted basic Equation (2), which quantifies the SS approximation, rewritten in the above notation, with $G' = G(1 - \Gamma)$:

$$[K/\dot{M} - 1]\ vdr/dr = G'\dot{M}/r^2 \qquad \text{(b) or KPP[2]}$$

indeed produces a gravity-like $v(r)$, by its integral from its lower boundary, r_i to r in the region bounded by: $r_i < r < r_x$:

$$v(r)^2 = v_i^2 + v_{esc}^2 \, [K/\dot{M} - 1]^{-1} \, [1 - r_i/r] \quad , \tag{c}$$

with

$$v_{esc}^2 = 2G'M/r_i \quad .$$

Thus, the particular values of L, $\dot{M}$, $L/\dot{M}$ affect *only* the particular values of: (1) the acceleration $v \, dv/dr$ in Equation (b), (2) the coefficient of the r-dependence of $v(r)$, and hence (3) the "asymptotic" velocity in Equation (c). But they do not affect the form of $v(r)$—as above, *this form is imposed by the approximation*. Contrary to the authors' assertion, the quantities (1)-(3) are not uniquely determined by the escape velocity, v_{esc}, in applying the boundary conditions, as we see below.

Also, calling the maximum velocity given by Equation (c), $v(r_\infty)$, the "asymptotic" velocity implicitly assumes that whatever regions may lie *exterior* to the SS (or S) region have no effect on $v(r)$. We return to this point in Section b, where we also discuss the relation between L and $\dot{M}$ implied by Equation (b) and the boundary conditions. We also remark that, by asserting that an $(L, \dot{M})$ relation and the relation $v(r)$ are uniquely determined by the SS-defined region, this hot-star theory resembles the original solar hot-coronal theory in presenting an algorithm to "predict" the size of the star's mass loss from only the properties of an outer region of the star's atmosphere. The regions below, and including, the transthermal region(s) are thus assumed to have no influence on $\dot{M}$ and $v(r)$. Of these lower regions, one demands only that the outflow velocity become negligibly small with depth. As in the solar case, one often hears in hot-star theory discussions, that conditions in the outer atmosphere "pull out" the mass from below—again missing the basic physics, especially the nonthermal aspects.

Because the SS (and S) approximations also exclude collisional source-sink terms in radiative transfer and dissipation, etc., they exclude all microscopic coupling between internal energy of the gas and that of the radiation field. The thermal structure of the gas in this atmospheric region defined by the SS and S approximations can, in consequence, only be assumed, not computed self-consistently. This leaves vague the physical location of the upper and lower boundaries of the region. Radiation transfer is restricted to including only scattering terms. Therefore, we note that the focal points in theoretical study of the coupling between mass outflow and atmospheric structure, as well as the non-LTE "soul" of the radiation diagnostics used in their observational study—which rest on these microscopic and macroscopic particle and photon interactions—are excluded in both SS and S approximations. Thus, one cannot hope to establish any relations between the mass outflow, including its dependence on radiative acceleration, and *the atmospheric structural pattern* involving RE, non-RE, superionized and subionized regions, etc., under these approximations.

At most, we may learn something about the relation between the imposed characteristics of this atmospheric region studied—*if* a region so defined actually exists—and that size mass flow entering the base of this region, which the radiation field in the region can accelerate into a mass loss. Note that, for this possibility to exist, the outflow velocity at the region's base must be subescape even though it is strongly superthermal, which places an upper limit on the T_e value there. Because the SS and S approximations also impose isothermality (without saying how to maintain it), T_e is bounded for the entire region.

Finally, from this same orientation, *we note the generally accepted belief that the mass outflow size is fixed by conditions below the thermal point; only its velocity is strongly affected by conditions in the superthermal regions. If this belief is correct, then—by studying the mass outflow only in the*

(SS or S) high-superthermal region—one cannot hope to establish that the characteristics of the radiation field in this region suffice to predict what mass outflow originates in the regions below it. One can only ask what kind of acceleration the radiation field in this region gives to a mass flow whose size has been specified by including in any study those regions below that to which the SS and S characteristics restrict us. That is, we must take *both* L and $\dot{M}$ as having specified values at the lower boundary of the SS or S region. Then we ask, via SS or S equations, what $v(r)$ that combination of L and $\dot{M}$ produce in this region. If that $v(r)$ reaches escape values, that $\dot{M}$ is "possible" for the star defined by that L and gravity. Whether the star can produce those $\dot{M}$ that satisfy this criterion cannot be decided from studying only the SS/S region. *Therefore, radiative acceleration, not radiative origin, is indeed the correct label for this theory.*

Thus, the overall atmospheric structure implicitly assumed *a priori* in SS and S modeling has four parts: (1) a lower atmosphere whose lowest regional structure is quasi-HE and RE, controlled by gas pressure and gravity, across which the outflow velocity evolves from some small subthermal value (origin unspecified) up to velocities where line radiative acceleration just becomes significant. Aside from diminution of the effective gravity by radiative acceleration in the continuum, the problems of understanding this region are common to Sun and hot-stars: compatability of $\dot{M}$ and v (HE) and their origin; (2) a transthermal range, in which radiative acceleration becomes increasingly important, and into the superthermal range, where gravity and radiative accelerations increasingly dominate; (3) the SS (or S) region; and (4) a region which the mass outflow enters at superescape velocities, but whose outer parts can differ significantly—in both thermal and macroscopic-velocity characteristics—from the lower regions.

Again we emphasize that, as for the Sun, the hot-star SS (or S) theory, and its algorithm used to "*predict*" the size of the mass loss, rest on considering only one, outer, region of all those constituting the whole atmosphere. And again similar to the solar case, that prediction focuses on the acceleration of a given-size mass outflow, incident on the base of the region, not on how, or what size, mass outflow actually originates in the underlying regions. We put this structure into perspective by making algebraically explicit the preceding remarks.

β. Solutions of the SS Equations. From the preceding discussion, we see that the specific objective in the solution of the equations—under the SS (or S) approximation—*should* be to show that there exists a relation between the character of the radiation field within, and the size of the mass outflow entering, the SS/S region which must be satisfied if that size mass outflow is to be further accelerated across the region to reach escape velocity (i.e., produce a mass flux from the star). A corollary result would be the conclusion that a steady-state mass outflow of some other size cannot be accelerated to escape velocity across this region, but must fall back into the star. If the latter were the *only* size mass outflow the star could produce, the star would remain a closed system. So as in the solar case, a star must satisfy two criteria to be an open system: that it can produce a mass outflow of some size; and that it can accelerate it to escape. Current SS (and S) theory says that the SS/S atmospheric region controls the latter criterion for hot stars.

For such a region to be a part of the atmosphere of a star from which a mass flux is observed: (1a) either *all* mass outflow values can be accelerated to escape velocity from that star so (1b) sub-SS regions produce whatever $\dot{M}$ they can, and the SS region accelerates it to escape, so $\dot{M}$ therefore tells us nothing of SS-region properties, only of sub-SS capacity to produce $\dot{M}$); or (2a) only *some* $\dot{M}$ can be so accelerated, and (2b) the sub-SS regions must be able to produce an $\dot{M}$ in that permitted range. Observations of the existence of an $\dot{M}$ in that permitted range should then tell us something about both SS and sub-SS regional properties. A most interesting aspect would be how and why SS and sub-SS regions should be able to mutually "adjust" in such ways as to show correlated behavior in producing a limited range of mass-loss values.

The Chapter 4 summary "pictures" *some* $\dot{M}$ = *one unique* $\dot{M}$ for a given value of (L,g). In fact, however, and contrary to what they profess to show, their SS approximation only establishes a range of possible values for the ratio of L to $\dot{M}$, not a unique value. And the r^{-2}-accelerated velocity law, existing because the SS/S approximations so prescribe it, admits a range of asymptotic velocities for a given value of v_{esc}, corresponding to the range in permitted values of the $L{:}\dot{M}$ ratio—again contrary to the Chapter 4 results. To show this, we return to Equations (b) and (c) and ask what $\dot{M}$ can be accelerated to escape velocity somewhere within the SS region. We set upper and lower limits on $\dot{M}$ as follows:

Upper-limit: A given radiation field can only accelerate some limited mass to escape. So, for the outward acceleration to be positive:

$$K/\dot{M} - 1 > 0, \text{ hence } \dot{M} < K \quad . \tag{d}$$

Lower-limit: For too-small $\dot{M}$, Equation (b) gives $dv/dr > 0$, but possibly too small to accelerate that $\dot{M}$ from v_i to the (local) escape velocity within the boundaries of that SS region of size limited by: $r \leqslant r_x$.

From Equation (c), we express this condition as:

$$v_{esc}{}^2\,(r = r_x) < v^2\,(r = r_x) = v_i^2 + (K/\dot{M} - 1)^{-1}\,(1 - r_i/r_x)\, v_{esc}^2\,(r_i) \quad . \tag{e}$$

We can estimate r_x/r_i from $V_{esc}\,(r_i)$ and q, via the second inequality defining the SS region and approximation—$4q^2 << 2G'M/r_x = v_{esc}{}^2(r_x)$—replacing it by the equality. Then, such a value for r_x is an upper limit on it. So we use:

$$r_i/r_x = v_{esc}{}^2\,(r_x)/v_{esc}{}^2(r_i) = 4q^2/V_{esc}^2\,(r_i) \tag{f}$$

to obtain from Equation (e):

$$K/\dot{M} < [V_{esc}^2\,(r_i) - v_i{}^2]/[4q^2 - v_i^2] \tag{g-1}$$

hence:

$$\dot{M} > K \cdot [4q^2 - v_i^2] \;/\; [V_{esc}^2\,(r_i) - v_i^2] \quad . \tag{g-2}$$

Thus, we have those values of an $\dot{M}$ which, if incident on the lower boundary of the region defined by the SS approximation, will be accelerated to the local escape velocity at some radius in the region bounded by $r_i < r < r_x$, with r_i/r_x defined by Equation (f):

$$K > \dot{M} > K\,[4q^2 - v_i^2] \;/\; [V_{esc}{}^2(r_i) - v_i^2] \quad . \tag{h}$$

These limits give all that can be said about the size of a mass outflow which can be accelerated to escape velocities across this low-temperature (to ensure v_i is subescape), noninteracting-particle, atmospheric region. Thus, from this SS approximation viewpoint, one would say that the theory *predicts* that the mass loss from the star should have a value lying within this range—but *not* a unique value fixed by *(L,g)*—without providing insight as to how and why the star produces such mass loss. In this sense, this "hot-star, cool mass-loss" theory is similar to the solar "cool-star, hot mass-loss" theory.

Introducing this range of $K/\dot{M}$ into Equation (c), we obtain the *range* of velocity distributions and of $v(\infty)$, the asymptotic velocity, that is compatible with a given $V_{esc}(r_i)$—*not unique values.*

We should clarify the difference between these results and those in Chapter 4, where the SS approximation leads its authors to predict unique values for $\dot{M}(L)$ and for $v_\infty(v_{esc})$. In Chapter 4, the authors define a quantity:

$$\epsilon = K^{-1} \dot{M} \, G'M \, (r^2 v \, dv/dr)^{-1} \quad , \tag{i}$$

and rewrite Equation (b) as:

$$\dot{M} = [1 - \epsilon] \, K = [1 - \epsilon] \, N_{eff} \, L/c^2 \tag{j}$$

to "picture" the single-valued relation between $\dot{M}$ and L. At various epochs, the authors have variously argued:

1. "An ϵ constant over the range of stars considered, demonstrates the proportionality of $\dot{M}$ to L; hence the predictability of $\dot{M}$ from L." Unfortunately, nothing in the SS approximation suggests such ϵ-constancy among stars. An estimate of the value of ϵ from its definition, Equation (i), requires a knowledge of $K/\dot{M}$. Substituting the definition of ϵ from Equation (i) into Equation (j) gives an identity in $\dot{M}$, which tells us nothing.

2. From Equation (j), an N_{eff} independent of r implies an ϵ independent of r; hence, a velocity distribution, neglecting v_i:

$$v^2(r/r_i) = (1-\epsilon)\epsilon^{-1} \, v_{esc}{}^2[1-r_i/r] \quad . \tag{k}$$

Above, we showed that such a velocity distribution is *imposed in the formulation of the SS approximation*, giving—without discussion—Equation (c), which is Equation (k) when ϵ is defined and substituted in Equation (c). ϵ is inherently constant in r. However, the authors further argue that observations show ϵ *to be constant across the sample of hot stars considered*: "to first order, it is independent of stellar parameters for luminous OB stars, with given metallicity." If this is true, one can draw conclusion (1) from Equation (j).

Unfortunately, the observational evidence shows the contrary to their conclusion. I abstract my 1983 summary: (1) In Chapter 4, the authors cite a 1978 study by Abbott as showing, to first approximation, $v_\infty \sim 3 \, v_{esc}$. Abbott discarded half of the 65 observed stars in the sample as not showing abrupt blue edges; hence, being unsuitable for testing the theory. A least-squares fit of a linear relation between v_∞ and v_{esc} gives an intercept of 540 km/s and a coefficient of 2.4. (2) Observational compilations by Hutchings and von Rudloff (1980), Cassinelli and Abbott (1981); reviews by Barlow (1982), Abbott (1982a, 1982b); suggest v_∞/v_{esc} lies near 3 only for mid-hot stars, decreasing on each side. Abbott's 1982a calculations give a quasi-constant ratio ~ 1; those of 1982b give 2.5 near T_{eff}

of 5.10^4 K, 3.5 at $2.5.10^4$ K, and 1 at 1.10^4 K. Thus, *neither observations nor some theoretical calculations of ϵ support the idea of its constancy among a large variety of stars.*

Therefore one can hardly use considerations on the form of $v(r)$ to support the idea of an isomorphic association between L and $\dot{M}$, much less that the value of L uniquely *determines* $\dot{M}$. Indeed, Abbott's latest summary (1985) states: "Little heed should be paid to the detailed shape of $v(r)$—there is insufficient physics in the model." His focus there lies on such things as the variation of the g_{rad} of Equation (a) with radius coming with changes in ionization, and the presence of other forces such as rotation and magnetic fields. These are legitimate concerns; because, as we emphasized above, neglect of all these is what produces *a priori* that gravity-like form of $v(r)$ stressed in Chapter 4—and indeed continually by Abbott himself, in spite of this quotation—as a major result of the theory. Apparently an ambivalence remains, a fascination with the significance of the observed values of v(max) and their *trend* to increase with brighter, more massive stars. Recall again the comments on the trend of v(max) across the HR diagram (shown in Figure 3-33) of Thomas, 1983. The problem is whether one observes atmospheric regions in a given star distant enough for v to have reached the true v(max), and whether it is time-variable. One can then use this information to ask the adequacy of theory in representing such observations. For added perspective, consider these two points: variability and "distant" regions.

First, note that Abbott reemphasizes what the Chapter 4 authors wished to show with their SS example: their belief that the theory is basically sound because the observed $\dot{M}$ follow the predicted scaling with L; and v_∞ with v_{esc}; it is only the coefficients which differ from the theory. However, the latter is the essential point, the situation is not linear, and these coefficients depend on $L/\dot{M}$ and V_{esc}, as Equations (c) and (h) show. The scaling *predicted* from, for example, the Chapter 4 misleading exposition (i.e., constant ϵ) of the SS theory is not the scaling actually given by self-consistent application of that theory. Specifically, the statement that $\dot{M}$ scales as L is only superficially valid. From Equation (h), one sees that both upper and lower limits on $\dot{M}$—hence also the intermediate values—are proportional to K, hence L. But the proportionality coefficient differs drastically: 1 for the upper limit; $\sim 4q^2/V_{esc}^{\,2}(r_i)$, neglecting v_i^2, for the lower. *A priori* one does not know where $\dot{M}$ lies in the range between upper and lower limits for a given star. As discussed, one must consider the sub-SS region to decide. Therefore, in asking the "scaling with L" between stars of different T_{eff} and V_{esc}, all three factors, as well as N_{eff}, enter; and the effect of the coefficient is hardly minor. For a 30000 to 40000 K star of 10 to 20 $M_\odot$, $R_\odot$, $4q^2/V_{esc}^{\,2} = o(100)$. Therefore, if one star at one epoch has $\dot{M}$ near its minimum value; and another star has $\dot{M}$ near its maximum value; this difference arising in the coefficient of L can compensate for a factor of 100 difference in L—*under the SS theory*. This "nonlinear" part of such L scaling cannot be ignored in assessing the applicability of the radiative-acceleration theory.

But this nonlinear scaling is not a liability of the radiative acceleration theory regarding observations, rather, it is an asset. First, it "interprets" the observational scatter about any linear relation. Equally important, however, it admits an observed variability to the theory's coverage. If the theory demands unique $(\dot{M},L)$ and (v_∞,v_{esc}) relations, any observed variability in $\dot{M},v_\infty$ requires (unobserved) variability in L and v_{esc}. Loosening such a uniqueness demand in accordance with the preceding SS-results increases the possible applicability of at least a modified version of such theory.

In his 1984 summary, Abbott further elaborates: "a legitimate worry is the relevance of any steady-state, homogeneous model for winds, whose basic instability is now clear from both observations and theory." Such "instability" has two aspects: that associated with variability, and that with heating (X-rays, for hot-star theoreticians). Consider the SS approach relative to variability. Self-consistently applied, it demonstrates only that, if radiation provides that acceleration which amplifies low-initial-velocity mass outflow into mass loss, the $(L,\dot{M})$ relation is only homomorphic, not isomorphic. However, this homomorphism gives a basis for showing that the radiative acceleration theory might be compatible with at least *some* kinds/amplitudes of variability.

One must accept that L fixes the $\dot{M}$ transmitted by the SS region, not produced by it; and that $\dot{M}$ has a range of values, not a unique value. (Whether the numerical range is changed by more complete theory than SS, and by N_{eff} differing from the SS-assumed constant value, is not important here; that it is a range of $\dot{M}$, rather than a sole $\dot{M}$, is.) Then, so long as the observed variability in $\dot{M}$ and $v(r)$ is encompassed by the "theoretical" range in what $\dot{M}$ can be transmitted by the SS region, the theory *may* be able to embrace that particular variability. The important aspect of such elasticity is the focus on transmitting/filtering via the radiative acceleration, rather than on originating, $\dot{M}$. The SS theory is compatible with "transmitting," not originating, some varieties of variability. If, of course, one shows that such transmitting/filtering is unstable, that process may also introduce a variability—this being the focus of attempts to produce X-rays in cold radiatively accelerated mass outflows. But one does not need an unstable superthermic mass outflow for SS (and more refined) theory to "live" with variability. Then, given a variability of mass inflow into the SS region falling in the range that can be transmitted by the region, one admits the possibility of variable $v(r)$, fixed by the values of ($\dot{M}$, L, V_{esc}, N_{eff}, etc.).

Then, *the observed variability in each of $\dot{M}$, v(r), etc. must arise either in the value of the mass outflow produced by regions below the SS-defined one, or in a variability of those properties of the SS region which fix the radiative acceleration.* From the preceding, we see that the latter can be only a variability in N_{eff}, which itself must come from variability in the local values of T_e, ρ, and opacity. Variability of T_e(local) requires variability in L—usually not observed—or in nonradiative heating, which requires its existence; we then reach the same result as that demanded by the presence of the instability associated with X-ray emission. Variability in ρ or opacity requires variability in $\dot{M}$, which, to avoid circular arguments, requires variability in mass outflow from the sub-SS regions. In any event, therefore, Abbott's worry on the relevance of steady-state wind theory to interpreting $v(r)$ reduces to a demand for some way to introduce either or both of variability in mass outflow from subatmosphere into the SS-region, and the existence + variability of nonradiative energy dissipation.

Therefore, abandoning the idea that the radiation field of a star uniquely determines the size of its mass loss; and that its gravity uniquely fixes the maximum value of the velocity of that mass outflow; a radiative acceleration of that mass loss *may* be compatible with observed variability in mass loss and its $v(r)$. Such reorientation of radiative acceleration thinking is not limited to the SS (or S) approach, we can simply again use them for illustration;—with no illusions on the numerical reliability of the results, as the Chapter 4 authors and Abbott's remarks have stressed.

Again realizing that SS (or S) theory tells us only about an $\dot{M}$ accelerated through the SS (or S) regions, we ask what it tells us on the size of that $\dot{M}$ and its variability. From Equation (h), the maximum size of $\dot{M}$ is $K = LN_{eff}/c^2$. For $N_{eff} = 1$, this is the Lucy-Solomon original prediction of the maximum size for $\dot{M}$. As they recognized, this is the size mass-loss, by thermal-nuclear burning, that produces the observed luminosity L. Indeed, we recognize L/c as the momentum given by completely absorbed radiation to accelerate mass loss $\dot{M}$ to escape at the velocity of light. Effectively, Equation (c) predicts $v_\infty = \infty = c$ for $\dot{M} = L/c^2$. Therefore, this $N_{eff} = 1$ solution is not useful; it is clearly a lower, rather than an upper, limit on the star's mass loss (nuclear-burning + aerodynamic outflow).

The physics of mass loss by this variety of aerodynamic outflow is thus contained wholly in the value of N_{eff}. Indeed, an upper limit on the outflow velocity should be c; and on the $\dot{M}$(max), it is K ($N_{eff} > 1$). Applying the same logic, $\dot{M}$(min) is $K \cdot 4q^2/V_{esc}{}^2(\mathrm{i})$, and the corresponding v_∞ is $2q = v_{esc}(r_x)$. As already recognized by the nature of this "free-particle" SS approach, $2q$ is indeed a minimum "escape" velocity, as in the original solar theory; in this thermal limit, there is effectively no radiative acceleration. Classical solar, and SS, theories coincide in the "weak mass-flux limit."

Clearly, the range permitted by these conditions—$2q$ to c for v_∞—and correspondingly, $(4q^2/V_{esc}^2)K$ to K for $\dot{M}$—is very broad. For a 30000K star, $4q^2/V_{esc}^2$ is o(100), so that is the mass-loss range for a given star. Because the actual value depends on N_{eff}, the physics is again contained in its computation.

Under this SS approximation, therefore, there are only three physical problems to clarify and solve: (1) the value of N_{eff}; (2) because a correct N_{eff} requires considering a variety of ionic species, and because their radiative accelerations will differ, one must address the problem of how noninteracting, differentially-accelerated species can be described by a "single-stream" velocity field; and (3) what produces, and the details of, the mass outflow from below the SS region into it. The existence of variability is no longer a problem under this "modified-SS" approximation. Its origin requires study of the sub-SS region. Thus far, theoretical work has essentially focused only on problem (1). Thus, the observed existence of variability, and what is required to produce a physically self-consistent representation of mass loss under SS theory, demand, alike, adjoining a consideration of the sub-SS region to that of the SS-region in order to provide proper lower boundary conditions for treating the SS region. Radiative acceleration theory and hot coronal theory have the same problem in this respect.

Turn now to the problem of outer boundary conditions on the SS region, hence the post-SS regions, which begin where $4q^2$ becomes significant relative to $V_{esc}^2(r)$. Continuing to use the SS approximation of Equation (a) for g_{rad}, Equation (b) becomes:

$$[K/\dot{M} - 1]\ v dv/dr = G'M/r^2 - 2q^2/r \quad , \tag{b-1}$$

and for sufficiently large r, the right-hand side (RHS) of the equation, hence dv/dr, becomes negative. For the acceleration to remain positive, K must decrease (i.e., N_{eff} must decrease). Castor, Abbott, and Klein long ago identified this r, where the RHS of Equation (b-1) would vanish because $V_{esc}(r)$ has decreased to $2q$, as the same "critical point" in radiative acceleration models as occurs in hot corona models. Here, mass outflow at $v(r_c) \geqslant q$, with $dv/dr\ (r > r_c) \geqslant O$, becomes mass loss under *any* theory. Therefore, either the left-hand side (LHS) of the equation also vanishes here and is negative for greater r, or the flow decelerates for $r > r_c$. The critical point is discussed at length by Abbott (1985), but neither it nor the post-SS regions are considered in Chapter 4. I try to put it into perspective. There are two possible effects of the existence of this post-SS region: (1) its perturbation of observations of the SS region to infer v_∞ and $\dot{M}$; and (2) its perturbation of SS theory via its perturbation of outer boundary conditions.

(*1*) If the preceding SS theory describes the SS region even grossly, we can effectively ignore the perturbation of the post-SS region on observations. From Equation (f), it does not begin until $r_c/r_i \sim V_{esc}^2(r_i)/4q^2 \sim$ o(100). By then, $v(r)$ is 99 percent of v_∞, and for constant q, the $2q^2/r$ term adds a term $-2q^2 \ln r/r_x$ to the $v(r)$ of Equation (c). The hot-star theoretical results give $v_\infty^2/2q^2 \gtrsim 10^3$ (cold wind); so there is no significant deceleration in atmospheric regions dense enough to observe. Therefore, v_∞ (SS) will be observed in the SS region, not perturbed by the post-SS regions, *if* SS theory is applicable. We note, from the discussion of the SS region, that larger v_∞ accompany larger $\dot{M}$. So because larger $\dot{M}$ give larger particle concentrations, it should be easier to observe the larger v_∞—again, *if* SS theory applies.

(*2*) The location of this critical point is the same for hot coronal or radiative acceleration theories—*for stars of the same gravity and T_e (atm)*. The basic difference between hot stars and Sun lies just in T_e, which is determined by the degree of nonradiative heating. For the real-Sun, r_x is 2 to 3 r_i;

for theoretical hot stars, it is o(100). Even if radiation provides the dominating hot-star wind acceleration, the value of T_e is critical for delineating regional structure, especially the extent of the SS region. The evolution of hot-star theory seems to be more focused on the effects of rotation, etc., on the location of the critical point than on effects of variability and heating; observations suggest otherwise.

Under hot coronal theory, mass outflow becomes mass loss at r_c if the outflow velocity reaches q there, to make the total velocity equal escape velocity. If it does not, the flow remains subthermal and decelerates. It cannot "recover" by accelerating again, as in the several "shock" decelerations in the underlying region where escape velocity exceeds thermal velocity. Therefore, satisfying conditions at the critical point depends directly on the flow velocity at r_c. $\dot{M}$ must be large enough to have produced large enough low-atmospheric flow velocity that the nonradiative acceleration brings the gas to thermal velocity at or before r_c. As earlier discussed, there is a range of possible $\dot{M}$ values above this minimal one, via the "imperfect" flows being shock-decelerated and then reaccelerated. Constructing a $v(r)$ by imposing the *perfect*, shock-free, nondissipative flow; and integrating inward from the escape point; imposes an unjustified sub-r_c atmospheric structure, and subatmospheric behavior; so risks error. In the "perfect," shock-free, configuration, the flow is everywhere subthermic upstream of the critical point—"sonic" signals can propagate upstream from the critical point; in configurations with shocks, the flow is partially superthermic—such signals are inhibited by the shock.

By contrast, the hot-star, cool-atmosphere, critical-point balance has no direct connection to the outflow velocity there. It rests on the r-dependence of K being of a form to give: $K/\dot{M} > 1$, $r < r_x$; $K/\dot{M} < 1$, $r > r_x$ (*if* the SS form of the governing equation is retained). It depends on the r-derivative of the acceleration, K, not on the velocity, at r_c. The outflow is already superescape at the critical point—by definition of those $\dot{M}$ transmitted by the SS region; and—for a cold wind—is effectively undecelerated until the interstellar medium. So, unless the basic physical picture of the SS approach—the study of hot radiative acceleration of a cool atmosphere—is changed, the critical point *should* have little impact on modeling how radiative acceleration in the SS region "controls" $\dot{M}$ and $v/(r)$. There should be the same range in values of those $\dot{M}$ which can be accelerated to escape velocity. This would be true if one began integration to construct $v(r)$ from the lowest regions and an $\dot{M}$ "injected" there. And, as repeatedly emphasized, there is no problem incorporating a variability of that "injected $\dot{M}$" into such radiative acceleration modeling *if* the radiative acceleration is either independent of (v, dv/dr) or varies as $v\, dv/dr$, as in Equation (a). Because the flow is everywhere superthermic upstream of the critical point throughout the SS region, any "signal" transmitted upstream must be radiative, not "acoustic," affecting the acceleration upstream only if critical-point opacities are large enough to produce significant back radiation. Only if T_e rises significantly over the "cold-flow" values does the critical point occur in high-density regions, and then one must ask what influence any coronal radiation field has (cf. the Be-star abstract).

However, on the stated-basis and to remove any aspect of deceleration, Castor, Abbott, and Klein introduced, and Abbott elaborated, their "singularity and regularity" conditions, from which they integrate inward and outward from the critical point. By studying N_{eff} in detail, they constructed a $K(r)$ which modifies the simple constant value of the SS Equation (a). With this, applying their singularity and regularity conditions, and iterating on a value of $\dot{M}$, those who apply their algorithms integrate inward and outward from the critical point to produce unique $\dot{M}$ and $v(r)$ for specified L and V_{esc}. In such a way, they eliminate deceleration above the critical point and retain their type scaling laws for $\dot{M}$ and v_∞, but remove that homomorphism between L and $\dot{M}$ which permits a variability in $\dot{M}$ and $v(r)$ without requiring L and/or V_{esc} to vary. They find such solutions to be unstable against large fluctuations in $v(r)$—(o(10^3 km/s) in size)—as remarked regarding their approach to producing X rays. It is known only to those who do such computations how uniquely they fix $\dot{M}$ and $v(r)$—i.e., the stability of such solutions, integrating inward from the critical point, against admitting sizeable ranges in these quantities. I illustrate this last point to conclude this summary abstract: (1) by Abbott's estimate of $\dot{M}$ from consideration of the critical point, using Equation (a)-type g_{rad}, and (2) by comment on the difference between SS and S (which rests on nonconstant K) approximations regarding "uniqueness" and the way it is reintroduced.

(*1*) Abbott (1985) adopts the condition that the electron-scattering opacity of the mass outflow gives $\tau = 1$ at the photospheric radius:

$$1 = \tau_e = \int \rho \sigma_e dr = C.\dot{M}.\int [r^2 v(r)]^{-1} dr \quad , \tag{l}$$

uses a g_{rad} of the form of Equation (a); and applies the Castor, Abbott, and Klein regularity conditions to obtain a velocity law, *apparently* implied to be valid in the range $r_{phot} < r < r_c$, with which he integrates Equation (l) to obtain:

$$\dot{M} v_\infty = (L/c)(2 - \Gamma)/\Gamma \tag{m}$$

to exhibit a scaling of $\dot{M}$ with L, but which also depends on V_{esc}.

If, on the contrary, one uses the Equation (c)-type velocity law usually cited as one of the significant results of the radiative acceleration theory—with uncertainty only on the coefficients—between r_i and $r_c > 100\, r_i$; and an HE, exponential in-Δr, ρ-distribution in the $\Delta r = r_i - r_{phot}$, negligible- or subthermic-velocity region, one obtains, adjoining the integral over the $[r_{phot} - r_i]$ region to Equation (l):

$$\dot{M}(K/\dot{M} - 1)^{1/2} (G'M)^{-1} = C_1 [1 - C_2\, \rho(r_i)\, H \exp\{\Delta r/H\}] \quad , \tag{n}$$

where H is the density scale height in the HE, immediate sub-SS, region. Equation (n) contains the condition that $v_i^2(K/\dot{M} - 1)/V_{esc}^2 << 1$.) Its application restricts the range of $\dot{M}$, for given L, permitted by Equation (h), but it hardly gives the scaling of Equation (m), if we evaluate v_∞ from Equation (c). And, until the SS treatment is extended to the sub-SS region, it is clear that the sub-SS contribution to τ obscures what $\Delta\tau$ can be used to establish the above limit, hence fix the RHS of Equations (m) and (n).

Physically, applying this opacity condition to "predict" the size of $\dot{M}$ misses the point illustrated by the Wolf-Rayet stars. The objective of the radiative acceleration theory is to predict what $\dot{M}$ will be accelerated to escape by given L. If we add a supplementary condition "*and* not obscure the HE, RE photosphere," it excludes those large $\dot{M}$ which L may accelerate to mass loss—but which give the star a Wolf-Rayet-type configuration (if one accepts that Wolf-Rayet stars are those with such large $\dot{M}$ that any HE, RE regions are obscured).

(*2*) The S versus SS configurations differ only in the forms adopted for g_{rad}; the other terms entering the governing equations remain the same. The change introduces the r-dependent K discussed above:

$$g_{rad} = (r^2 v\, dv/dr)^\alpha M^{-\alpha} r^{-2} C' \quad , \tag{o}$$

where:

$$C' = (LN_o/c^2)\, q^{1-\alpha} C_1 \quad . \tag{p}$$

The logic of this form for g_{rad} is summarized in Chapter 4. In essence, the SS g_{rad} is based on only strong (i.e., saturated) spectral lines accelerating the flow. If one adds weak lines, their effect resembles that of the continuum: no $v \cdot dv/dr$ dependence to desaturate the absorbing region. N_0 represents the total of all lines contributing to the acceleration, weak and strong. The Castor, Abbott, and Klein approach treats lines of all strengths together, in producing the single-term expression, Equation (o), for g_{rad}. The parameter α simply represents an empirical fitting of this statistical combination of the weak and strong lines.

Clearly, this approach is open to question. If one simply introduced all strong lines in the form of Equation (a), and all weak lines as a modification of the parameter Γ, then the combined result would be that of the SS treatment, simply with differing values of N_{eff} and G'. That is, one would obtain the SS result of a range in $\dot{M}$ and $v(r)$—no unique $(\dot{M},L)$ and (v_∞, V_{esc}) relations—being accelerated to escape (or transmitted) by the hot-star atmosphere. By contrast, if one imposes the (v, dv/dr)-dependence of g_{rad} given by Equation (o) and takes literally a constant α (i.e., α independent of r), then one obtains unique values of $\dot{M}$ and $v(r)$, expressed by:

$$v dv/dr = \alpha(1 - \alpha)^{-1} \, G'M/r^2 \tag{q}$$

in place of Equation (b)—i.e., $v_\infty{}^2$ is uniquely $V_{esc}{}^2.\alpha/(1-\alpha)$—and a corresponding unique scaling of $\dot{M}$ with L via:

$$\dot{M} \sim L^{1/\alpha} \, \alpha/(1-\alpha)^{1 - 1/\alpha} \quad , \tag{r}$$

as shown in Chapter 4.

Kudritzki, Pauldrach, and Puls stress that α is not a free parameter, but that its value is fixed by detailed calculations of conditions in the wind solutions. This implies iteration between parameters of the radiative acceleration and the parameters of the resulting solution—and an r-dependence of α. Given that the solutions result from iterative integration inward and outward from the critical point, using a set of singularity and regularity conditions formulated in terms of the r-dependence of g_{rad} at the critical point—and the distant, low-density location of the critical point—it is not at all clear how reliable these solutions are in establishing uniqueness in values of $\dot{M}$ and $v(r)$. We again note Abbott's cautions on reliability of $v(r)$, even though he expresses confidence in the $(\dot{M}, L)$ results. One cannot but be impressed by the results that Kudritzki et al. summarize in Chapter 4, although the model-dependent nature of "observational results" on $\dot{M}$ and v_∞ again suggest caution.

Finally, the increasing evidence for variability and nonradiative heating summarized here leaves one unquiet with a modeling which continually focuses on establishing that "uniqueness" which makes it difficult to include variability in the theory. I prefer the SS modeling, subject to modification which guides its flexibility toward nonuniqueness, which the authors proposed to put the problems into perspective. They appear to accomplish their objective.

Richard N. Thomas
Boulder, Paris, 1988

ACKNOWLEDGMENT

The editors and authors wish to acknowledge the great contributions of Leo Goldberg to stellar astronomy and general astrophysics, both as a research scientist and as a major force in planning and managing important elements of the American programs in observational astronomy at ground-based observatories and in space. We were fortunate to have him as the Senior Advisor to NASA for this monograph series, both for his comprehensive grasp of American astronomy and as an internationally oriented scientific statesman, in which capacity, among his many distinguished roles, Leo served as the President of the International Astronomical Union during the period from 1973 to 1976. With his passing in November of 1987, we have lost a wise, firm counselor, and an ever encouraging friend.

The Editors and Authors
April 1988

CONTENTS

PART I

Introduction

Lucienne Divan and Marie-Louise Burnichon-Prévot

PART II

One Perspective on O, Of, and Wolf-Rayet Stars
Emphasizing Winds and Mass Loss, With Remarks
on Environment and Evolution

RÉSUMÉ

Première Partie

Dans la première partie de ce livre, rédigée avant 1984, on donne les informations de base sur les étoiles O et Wolf-Rayet, en indiquant comment elles sont définies et quelles sont leurs principales propriétés parmi celles qui sont dites observables. La frontière entre les propriétés "observables" et les autres est évidemment très arbitraire et dépend essentiellement de l'idée personnelle que l'on se fait sur la valeur des théories toujours nécessaires pour interpréter un phénomène quelconque. On considère depuis longtemps que la longueur d'onde d'une raie spectrale est une grandeur observable et nous avons admis que la température effective d'une étoile O en était une également, depuis que les mesures de flux se sont étendues à l'ultraviolet spatial. Par contre, nous avons considéré que l'âge d'une étoile relevait encore de la théorie et les problèmes liés à l'évolution stellaire ne seront pas discutés dans cette partie du livre relative aux observations.

Les étoiles O sont définies par les caractères de leur spectre dans le domaine visible: distribution d'énergie très bleue et présence d'absorptions dues à l'hélium ionisé. Elles ne forment cependant pas un groupe homogène et des étoiles ayant des spectres analogues peuvent avoir des masses, des rayons, des âges et des structures internes très différents. La plupart des étoiles O connues actuellement sont brillantes, massives, concentrées près du plan galactique et correspondent aux premiers stades de l'évolution stellaire. Ce sont les étoiles dites "O normales". Mais on découvre de plus en plus d'étoiles O peu lumineuses et peu massives dans le halo galactique. Ce sont les "sous-naines O". Leur structure interne est encore mal élucidée mais il est certain que ces étoiles chaudes et peu massives ne peuvent correspondre qu'à des stades avancés de l'évolution stellaire.

Le spectre des étoiles de Wolf-Rayet est dominé par des raies d'émission fortes et larges, dues à des éléments fortement ionisés.

Ces définitions étant données, on décrit les résultats obtenus en ce qui concerne la classification spectrale, les couleurs intrinsèques, les températures effectives, les masses et les rayons, ainsi que les principaux travaux qui ont conduit à ces résultats. Dans tous les cas, les étoiles O normales, les sous-naines O et les étoiles de Wolf-Rayet posent des problèmes entièrement différents et ces trois catégories d'étoiles chaudes sont traitées séparément.

Classification Spectrale. On indique que les premiers systèmes de classification ont été décrits en détail par Curtis (1932) et que les classifications développées avant 1965 sont discutées par Underhill (1966). Trois systèmes de classification concernant les étoiles O normales sont ensuite décrits: la classification de Walborn, celle de Conti et le système BCD. La classification de Walborn (1971a, 1972, 1973a) est un raffinement de la classification MK grâce à des spectres plus dispersés (62 ou 63 Å/mm). La table I-1 donne les principaux critères utilisés par Walborn et la table I-2 les étoiles standard définissant chaque type spectral. La classification de Conti est basée sur la mesure du rapport des largeurs équivalentes de deux raies, l'une de l'hélium neutre, l'autre de l'hélium ionisé sur des spectres à 16 Å/mm et permet de définir quantitativement les différents sous-types O (voir Table I-3). La classe de luminosité est indiquée par le rapport SiIV 4089/HeI 4143. Les types spectraux ainsi déterminés

sont donnés pour 150 étoiles O dans Conti and Leep (1974) et Conti and Frost (1977). La classification BCD est un système quantitatif basé uniquement sur des paramètres liés au spectre continu de l'étoile. Ces paramètres, qui sont la grandeur D et la position λ_1 de la discontinuité de Balmer (voir la figure I-1 pour la définition de λ_1) dépendent directement de la température effective et de la valeur de la gravité qui caractérisent la photosphère de l'étoile. Pour les étoiles O normales dont il est question ici, λ_1 et D permettent de caractériser la magnitude absolue de l'étoile et la figure I-2 donne le résultat de la calibration empirique du diagramme λ_1D en magnitudes absolues. La figure I-2 montre également la corrélation entre la classification λ_1D et la classification MK.

Une brève discussion du problème de la classification spectrale pour les sous-naines O est ensuite donnée et l'on indique l'impossibilité d'une classification à deux paramètres, telle que celle qui est pratiquement suffisante pour les étoiles O normales.

Le problème de la classification des étoiles de Wolf-Rayet est entièrement différent de celui des étoiles O normales par suite de l'impossibilité d'obtenir des critères basés sur une photosphère observable et que l'on puisse modéliser en termes de température effective et de gravité. Les critères utilisables pour classer les étoiles de Wolf-Rayet sont liés aux raies d'émission formées en dehors de la photosphère. On peut classer les spectres en formant des séquences le long desquelles l'aspect du spectre varie de façon progressive, mais rien ne permet de croire que, comme dans le cas des étoiles O normales, ces séquences correspondent à une variation continue et monotone de la température effective ou de la gravité. Le fait que dans un même objet on puisse observer des raies correspondant à des niveaux d'excitation et d'ionisation très différents indique bien la complexité du problème. De plus il n'y a pas pratiquement deux étoiles de Wolf-Rayet ayant des spectres réellement identiques. Les classifications qui sont décrites plus loin correspondent en fait à une description empirique des spectres.

La classification développée par Beals (1938) à Victoria sépare les Wolf-Rayet en deux séquences qui semblent plus ou moins parallèles en ce qui concerne la température: les WN5, WN6, WN7 et les WC6, WC7, WC8 dans lesquelles les émissions dominantes sont respectivement H, HeI, HeII, NIII, NIV, NV et H, HeI, HeII, CII, CIII, CIV, OIII, OIV, OV. Le problème de la relation entre ces séquences et la composition chimique de l'atmosphère de l'étoile est brièvement discuté. La classification des étoiles de Wolf-Rayet a été reprise par Hiltner et Schild (1966) puis par L. F. Smith (1968a, 1968b) dont les critères de classification sont donnés dans les tables I-4 et I-5. Van der Hucht et al. (1981) ont appliqué le système de Smith à un plus grand nombre d'étoiles et introduit de nouveaux sous-types.

Toutes les classifications décrites jusqu'ici étaient faites à partir de spectres visibles. Après une description des principaux programmes ultraviolets (OAO-2, OAO-3, TD1 + S2/68, TD1 + S59, Skylab S-019, ANS, IUE), on donne des indications sur les classifications faites à partir de spectres ultraviolets à faible résolution et sur les projets pour le futur.

Couleurs Intrinsèques. Toutes les étoiles O, sans exception probablement, sont affectées par l'extinction interstellaire et leurs couleurs intrinsèques ne sont pas observables directement. Le problème encore mal résolu de la détermination des couleurs intrinsèques est discuté, puis on passe en revue les différents moyens utilisés pour corriger les couleurs observées du rougissement produit par l'extinction interstellaire dans le domaine des longueurs d'onde observables depuis le sol: extrapolation des résultats obtenus pour les étoiles B (mais la base théorique de cette extrapolation est discutable), méthode des amas (l'hypothèse que toutes les étoiles de l'amas sont également affectées par l'extinction interstellaire n'étant vraie, au mieux, que statistiquement, cette méthode nécessite l'étude d'un grand nombre d'étoiles dans un grand nombre d'amas et sa précision est limitée), méthode des systèmes multiples (il est alors beaucoup plus facile d'admettre que l'extinction interstellaire est la même pour toutes les étoiles du système, mais le nombre des systèmes multiples étudiés en détail est encore insuffisant).

Les couleurs intrinsèques des étoiles O normales sont ensuite données dans trois systèmes photométriques (Table I-8 pour le système UBVRIJKLM, relation I-4 et Table I-9 pour le système uvby, Table I-10 pour le système BCD), ainsi que des indications sur la manière dont ces valeurs intrinsèques ont été obtenues. Enfin, la figure I-10 montre comment sont placées les bandes passantes des systèmes UBV et uvby par rapport au spectre d'une étoile.

Beaucoup de sous-naines O sont à de faibles distances et à haute latitude galactique; elles ne sont alors pratiquement pas affectées par l'extinction interstellaire et leurs couleurs intrinsèques sont observables directement. Ces couleurs sont généralement plus bleues que celles des étoiles O normales les plus bleues et des valeurs $B-V = -0.34$ sont couramment observées.

Les étoiles de Wolf-Rayet sont, comme les étoiles O normales, situées à de grandes distances et voisines du plan galactique. Toutes sont donc affectées par une extinction interstellaire dont on a cherché à s'affranchir par différentes méthodes. Les principales parmi celles utilisées entre 1966 et 1984 sont décrites. On est généralement obligé d'admettre, sans que cela soit réellement prouvé, que l'on peut observer entre les larges raies d'émission de l'étoile de Wolf-Rayet un continu qui viendrait d'une photosphère que l'on compare alors à des photosphères d'étoiles normales ou à des modèles, convenablement choisis on l'espère, pour en déduire l'excès de couleur de l'étoile (E(B-V) généralement) puis dérougir tout le spectre avec une loi d'extinction interstellaire connue. On a pu cependant s'affranchir de toutes ces hypothèses pour dérougir certaines étoiles de Wolf-Rayet, celles par exemple qui sont situées dans les Nuages de Magellan où l'extinction interstellaire moyenne est faible. Par contre, les essais faits pour dérougir les distributions d'énergie des étoiles de Wolf-Rayet à l'aide de la bande à 2200 Å se sont révélés décevants. Dans tous les cas, l'interprétation des résultats est rendue très difficile par le fait que les raies d'émission ont des intensités très variables même à l'intérieur d'un même sous-type Wolf-Rayet et que s'il y a un continu photosphérique observable entre les raies il est très difficile de l'obtenir même avec des résolutions élevées. Ceci explique les résultats plus ou moins contradictoires obtenus par différentes méthodes. Il reste que les étoiles de Wolf-Rayet ont des températures effectives élevées correspondant en moyenne à des types O9 ou B0. Aucune corrélation entre la température effective et le sous-type Wolf-Rayet n'a été confirmée.

Cette étude du domaine des longueurs d'onde observables depuis le sol est complétée par quelques indications sur deux sources de distributions absolues d'énergie, le catalogue de Breger qui est une compilation de distributions non corrigées de l'extinction interstellaire et une liste publiée par Kuan et Kuhi où les distributions d'énergie sont corrigées de l'extinction interstellaire.

Dans le domaine de l'utraviolet spatial, l'obtention des couleurs intrinsèques nécessite, comme dans le domaine visible, la connaissance de la forme de la loi d'extinction interstellaire et celle de la quantité de matière absorbante. La forme de la loi d'extinction a été étudiée par différents auteurs. Cette forme semble relativement constante (en ce qui concerne la matière interstellaire diffuse) depuis l'infrarouge jusque vers 2500 Å, mais présente des variations parfois notables d'un point à l'autre de la Galaxie pour les longueurs d'onde plus courtes. Les couleurs intrinsèques des étoiles O ont dû cependant être obtenues avec une loi moyenne, car la loi particulière éventuellement applicable à chacune d'elles est inconnue. La quantité de matière interstellaire admise pour chaque étoile est déduite de la valeur E(B-V) obtenue par l'étude du domaine visible, faute d'autre moyen pour la déterminer.

Les principales sources de distributions d'énergie ultraviolettes sont recensées et six différentes grilles de couleurs intrinsèques, obtenues en corrigeant les distributions d'énergies observées de l'extinction interstellaire, sont comparées entre elles (Figure I-12) et à des modèles stellaires. Chacune de ces grilles correspond à une origine et/ou un traitement différent des données: nombre des résultats pris en considération, méthodes de lissage, origine de la calibration absolue... L'accord entre les grilles différentes est assez bon pour les longueurs d'onde supérieures à 1600 Å, moins bon pour les longueurs d'onde plus courtes. L'accord avec les modèles stellaires dépend de la température effective que l'on adopte pour le type spectral considéré. Avec les échelles habituelles, les étoiles chaudes, surtout si elles sont supergéantes, sont plus rouges que les modèles. Indépendamment de toute référence à une échelle de température effective, on peut faire coïncider un modèle de Kurucz (1979) avec une distribution intrinsèque jusqu'à environ 2000 Å, mais pour les longueurs d'onde plus courtes le modèle devient progressivement trop bleu.

Températures Effectives et Corrections Bolométriques. Trois méthodes sont décrites pour déterminer les températures effectives des étoiles O normales. La première méthode consiste à comparer les largeurs équivalentes ou les profils observés pour les raies spectrales à ceux prédits par des modèles d'atmosphère. Les échelles de températures effectives correspondant aux types spectraux O4 à O9, présentées sur la figure I-13 pour les classes de luminosité V et I (ou f), ont été obtenues par Conti à partir du rapport des largeurs équivalentes de HeI λ4471 et HeII λ4542. Dans la table I-12, on donne les températures effectives et les gravités que Kudritzki et ses collaborateurs ont obtenues pour cinq étoiles O en comparant les profils observés à grande dispersion pour toute une série de raies de l'hydrogène et de l'hélium à ceux prédits par les modèles stellaires très élaborés qu'ils avaient construits.

Dans la deuxième méthode, on détermine les températures effectives en comparant la distribution de l'énergie dans les flux stellaires à celle de modèles d'atmosphère. Pour les très hautes températures correspondant aux étoiles O, le domaine spectral directement observable au sol est trop restreint pour que la méthode puisse donner des résultats acceptables mais, dans le cas des étoiles O excitatrices de régions HII, on a pu évaluer le flux stellaire dans le continu de Lyman par la méthode de Zanstra et déterminer la température effective par comparaison au flux observé dans le domaine des magnitudes visibles. La figure I-13 et la table I-12 donnent les résultats obtenus ainsi par Morton.

La troisième méthode, la plus directe, a pu être utilisée pour les étoiles O depuis que les observations par satellite ont permis d'atteindre les régions du spectre où ces étoiles émettent la majeure partie de leur rayonnement. La partie observable du rayonnement est corrigée de l'extinction interstellaire et la partie non encore observable (qui peut atteindre 50% du flux total dans le cas des étoiles les plus chaudes) est déduite des modèles stellaires, soit en totalité soit partiellement, le flux correspondant au continu de Lyman pouvant être évalué par la méthode de Zanstra pour les étoiles excitatrices de régions HII. On obtient ainsi le flux total reçu de l'étoile et l'on montre facilement que ce flux permet de calculer la température effective de l'étoile si l'on connaît son diamètre angulaire.

Les étoiles les plus proches ont des diamètres angulaires assez grands pour être mesurables directement par les méthodes interférométriques, mais le nombre de ces étoiles proches est très petit et la plupart des diamètres angulaires ont été obtenus par des méthodes plus indirectes. Ces méthodes font intervenir des modèles stellaires, mais seules les régions spectrales (proche infrarouge en général) où ces modèles sont le moins sujets à caution sont utilisées pour la détermination des diamètres angulaires.

Les principaux résultats obtenus sont donnés dans la figure I-13 et la table I-12 où les températures effectives sont présentées en fonction du type spectral. Pour les types O les plus avancés, la concordance entre les résultats donnés par les diverses méthodes est relativement bonne, avec cependant des températures un peu plus basses pour les échelles faisant intervenir la méthode de Zanstra. On observe une relation entre la température effective et le type spectral, avec des températures plus basses pour les supergéantes que pour les naines à l'intérieur d'un même sous-type. Mais pour les types O5 et moins avancés, les températures effectives calculées à partir des flux intégrés ont une dispersion très grande à l'intérieur d'un même sous-type O, indiquant probablement que les raies définissant le type spectral se forment dans des conditions très différentes de celles qui règnent à l'endroit où le spectre continu se forme.

Pour les sous-naines O, on ne peut obtenir les températures effectives directement à partir des flux intégrés et des diamètres angulaires. En effet, les diamètres angulaires sont tous encore largement inaccessibles aux mesures interférométriques et leur évaluation à l'aide de modèles d'une atmosphère dont on ignore a priori les paramètres de gravité et de composition chimique n'a pas été tentée. De plus, les températures sont si élevées que la partie non observable du flux, qui est difficile à évaluer à partir de modèles pour les raisons données plus haut, est souvent plus importante que la partie observable.

Les températures effectives ont donc dû être obtenues de façon indirecte, et diverses méthodes d'abord sommaires puis plus élaborées ont été utilisées entre 1966 et 1975. Finalement, Kudritzki a calculé à partir de 1976 une série complète de modèles d'atmosphère comportant toutes les valeurs de la température et de la gravité qui peuvent se rencontrer dans les sous-naines O et toutes les compositions chimiques depuis l'hélium pur jusqu'à l'hydrogène pur. L'étude de chaque étoile comporte alors l'observation d'un maximum de caractères dans le spectre (distribution de l'énergie dans le continu, discontinuité de Balmer, largeur équivalente et profils des raies de H, HeI, HeII...) puis le choix du modèle qui représente le mieux ces caractères. On obtient ainsi à la fois la température effective, la gravité et la composition chimique. Les résultats obtenus ainsi par Kudritzki et ses collaborateurs sont d'autant plus crédibles que le nombre des paramètres observés est très supérieur au nombre de paramètres de la grille de modèles et que la solution obtenue avec quelques-unes seulement des observations permet de représenter parfaitement toutes les autres.

Les résultats obtenus entre 1978 et 1982 sont donnés dans la table I-13 et à titre de comparaison on donne dans la table I-14 les mêmes paramètres obtenus par les mêmes méthodes pour une série d'étoiles centrales de nébuleuses planétaires.

Les températures effectives des étoiles de Wolf-Rayet ont été déterminées en appliquant les mêmes méthodes que pour les étoiles O normales mais des difficultés supplémentaires assez considérables viennent de la présence de larges raies d'émission dans tout le spectre et d'un continu non thermique dans l'infrarouge qui s'étend peut-être même dans le domaine visible. Les raies d'émission gênent dans la détermination du continu et surtout rendent l'évaluation de l'extinction interstellaire très incertaine. Le continu non thermique, bien que l'on tente d'en corriger les effets, peut fausser la détermination des diamètres angulaires.

Les principaux travaux publiés entre 1970 et 1982 sont décrits et les résultats donnés dans la table I-15. La concordance entre les valeurs obtenues pour une même étoile par des méthodes différentes est relativement bonne. La plupart des températures effectives se situent dans la fourchette 25 000 - 35 000 K et les Wolf-Rayet ne semblent donc pas être des étoiles particulièrement chaudes. On n'observe pas en général de relation entre la température effective et la séquence ou le sous-type de l'étoile; cependant il semble que les WC5 et les WN3, 4 et 5 aient des températures plus élevées que les autres.

Pour terminer, on donne quelques indications sur les corrections bolométriques qui sont évaluées à partir de modèles stellaires. Les résultats correspondant aux étoiles O normales et aux sous-naines sont donnés dans la table I-16.

Magnitudes Absolues. Pour obtenir la luminosité absolue d'une étoile, il faut connaître la luminosité apparente que l'on observerait en l'absence d'extinction interstellaire et la distance. Les étoiles O normales, en très petit nombre par unité de volume et par ailleurs situées dans le plan galactique, sont toutes trop éloignées pour avoir une parallaxe trigonométrique mesurable et leur luminosité apparente est toujours affectée par une extinction interstellaire inconnue a priori. On doit donc, en fait, déduire leur magnitude absolue de celle d'autres types d'étoiles. On utilise pour cela les étoiles O appartenant à des groupes stellaires (systèmes multiples, amas, associations) qui contiennent des étoiles de magnitude absolue connue, étoiles B de la série principale en général, et on obtient les magnitudes absolues des étoiles O différentiellement en admettant que les effets de distance et d'extinction interstellaire sont les mêmes pour toutes les étoiles d'un même groupe. Les difficultés sont dues au fait qu'il est parfois difficile de décider de l'appartenance d'une étoile donnée à un groupe et que l'uniformité de l'extinction interstellaire à l'intérieur du groupe n'est qu'une première approximation. Dans le cas des systèmes multiples, ces difficultés sont beaucoup moins sérieuses que dans le cas des amas ou des associations, mais le nombre d'étoiles O étudiées est encore trop petit.

Les résultats obtenus montrent que la magnitude absolue est en relation avec le type spectral et la classe de luminosité, les étoiles classées "f" étant les plus lumineuses. Les résultats donnés dans la table I-21 sont les plus récents et proviennent de l'étude de 248 étoiles O dans 63 amas.

Toujours à l'aide d'étoiles situées dans des amas, on a obtenu également diverses formules empiriques permettant de relier la magnitude absolue à la largeur équivalente de la raie H_γ. Ces formules valables principalement pour les étoiles B ont été étendues aux étoiles O5 à O9 de luminosité V et O9 de luminosité III.

La présence dans la Galaxie d'une population d'étoiles bleues et très bleues, plus lumineuses intrinsèquement que les naines blanches mais beaucoup moins que les étoiles bleues de la série principale, a été démontrée pour la première fois en 1947 par Humason et Zwicky. Les étoiles les plus bleues de cette population ont des couleurs et des spectres analogues à ceux des étoiles O normales; elles forment le groupe des sous-naines O. Beaucoup sont situées en dehors de plan galactique et à des distances relativement faibles; l'effet de l'extinction interstellaire sur leur magnitude apparente est ainsi souvent négligeable. Par contre, aucune parallaxe trigonométrique n'est connue et les sous-naines O ne faisant pas partie comme les étoiles O normales de groupes stellaires dont on peut connaître par ailleurs la distance, il est difficile d'obtenir des informations purement empiriques sur leur magnitude absolue. De telles informations peuvent cependant être obtenues, statistiquement, par l'analyse des mouvements propres qui sont faibles mais non négligeables. Par ailleurs, dans un certain nombre de cas on a pu observer une sous-naine O dans un spectre composite et déduire la magnitude absolue de la sous-naine de celle de son compagnon. La présence d'une sous-naine dans un amas globulaire a pu être démontrée dans quelques cas et la magnitude absolue déduite de la distance de l'amas. On peut encore déterminer la distance de quelques sous-naines situées dans le plan galactique à partir de leur rougissement et de la distribution de la matière interstellaire dans leur direction. Le même travail a été fait aussi à partir des raies interstellaires. Enfin, la détermination spectroscopique des magnitudes absolues (à partir de la formule $M_V = 2.5 \log g + 10 \log \theta - 2.5 \log M/M_\odot - 5.82 - BC$) a pu être faite sur un nombre relativement grand d'étoiles mais elle ne peut être considérée comme purement empirique car on doit faire des hypothèses sur la masse de l'étoile. Le résultat est que dans le plan log g, θ les sous-naines forment à $\theta = 40.000$ K une séquence verticale s'étendant sur tout l'intervalle qui sépare les naines blanches des étoiles O de la séquence principale. Les magnitudes absolues vont de $-1,8$ à $+6,8$ mag, avec une valeur moyenne ($+4,3$ mag pour l'ensemble des résultats les plus faibles) comparable à ce que l'on obtient par les autres méthodes. Cependant, divers indices suggèrent que les sous-naines O ne forment pas un groupe homogène.

Les étoiles de Wolf-Rayet ont des magnitudes absolues très difficiles à déterminer car leurs distances sont trop grandes pour que les parallaxes trigonométriques soient mesurables et la réalité de leur association avec les régions HII, sur lesquelles elles se projettent plus souvent que le simple hasard ne le voudrait, toujours très difficile à établir réellement dans chaque cas particulier. Elles sont par contre faciles à détecter à grande distance et on en connaît une centaine dans le Grand Nuage de Magellan dont la distance est relativement bien connue et où l'extinction interstellaire est faible; mais certains sous-types Wolf-Rayet sont absents du Grand Nuage et pour les autres, bien qu'aucune différence spectroscopique ne le suggère, on peut se demander s'ils correspondent bien à des objets identiques dans le Grand Nuage et dans la Galaxie. C'est pourquoi les deux méthodes (étoiles du Grand Nuage et étoiles liées à des régions HII) ont été utilisées, conjointement ou non, par différents auteurs pour déterminer les magnitudes absolues des étoiles de Wolf-Rayet. La table I-25 donne les résultats des principaux travaux. La luminosité trouvée pour un même sous-type Wolf-Rayet est en moyenne plus faible dans la Galaxie que dans le Grand Nuage, mais les incertitudes sur les distances d'une part, et la dispersion des magnitudes à l'intérieur d'un même sous-type d'autre part sont telles que la réalité du phénomène est très loin d'être établie. La dernière version de la relation entre sous-type Wolf-Rayet et magnitude absolue est donnée dans la table I-26; elle tient compte d'étoiles Wolf-Rayet découvertes dans le Petit Nuage de Magellan, de l'étude de la multiplicité des étoiles du Grand Nuage et de la définition de nouveaux sous-types.

Dans le cas des étoiles de Wolf-Rayet, la magnitude absolue visuelle n'est pas définie dans le large domaine spectral habituel mais dans le domaine beaucoup plus restreint du filtre v de Westerlund de façon à éliminer, en partie tout au moins, l'influence de fortes raies d'émission très variable d'un sous-type à l'autre.

Masses et Rayons. Les données empiriques sur les masses et les rayons des étoiles O proviennent de l'étude des binaires spectroscopiques présentant deux systèmes de raies mesurables et des éclipses. En l'absence d'éclipse, on peut encore obtenir une valeur minimum pour les masses, m $\sin^3 i$, ou si le spectre de la composante secondaire est invisible, la valeur de la fonction de masse. Les résultats obtenus sont présentés dans la table I-28. Les déterminations complètes des masses et des rayons n'ont pu être faites que pour 7 systèmes et dans trois cas seulement les résultats peuvent être considérés comme ayant une bonne précision. Pour 21 systèmes les valeurs de m $\sin^3 i$ ont pu être obtenues et pour 12 autres les valeurs de f(m). L'examen de toutes ces données montre que la masse moyenne en unités solaires des étoiles O se situe aux environs de 25 et qu'aucune masse n'est très inférieure à 20. La discussion du système très complexe HD 47129 montre que les étoiles O peuvent avoir des masses très grandes, 50 ou même 100 masses solaires. Les rayons sont compris entre 4 et 19 unités solaires, avec une valeur moyenne voisine de 10.

Le nombre très faible des masses et des rayons que l'on a pu obtenir par l'étude des binaires a conduit à utiliser d'autres méthodes, beaucoup plus indirectes mais applicables à un grand nombre d'objets.

Pour déterminer les masses des étoiles O, on peut les placer dans un diagramme HR sur lequel on a tracé les trajets évolutifs d'étoiles de différentes masses. On obtient ainsi des valeurs comprises entre 20 et 100 masses solaires.

Les diamètres angulaires de quelques étoiles O ont pu être obtenus par des méthodes interférométriques ou par l'observation des occultations dues à la Lune. Des diamètres angulaires beaucoup plus nombreux ont été obtenus par une méthode indirecte, en comparant le flux observé (provenant de l'ensemble du disque stellaire) à celui d'un modèle convenablement choisi et qui donne le flux émis par unité de surface. Ces diamètres angulaires conduisent aux rayons linéaires lorsque l'on peut évaluer les distances, ce qui est le cas pour les étoiles O nombreuses appartenant à des amas ou des systèmes multiples. Les rayons peuvent également être déduits de la position de l'étoile dans le diagramme HR théorique.

Les résultats de ces méthodes indirectes, donnés dans les tables I-29 et I-30, forment un ensemble cohérent. Les valeurs des masses et des rayons donnés par les méthodes directes sont confirmées, leur variation en fonction du sous-type spectral est precisée.

Les modèles de structure interne suggèrent que les masses des sous-naines O sont, en unités solaires, voisines de 0,5 et que des masses nettement supérieures à 1 sont exclues, mais les données empiriques ou semi-empiriques sur ces masses sont si réduites qu'elles ne permettent guère d'infirmer ou de confirmer ce résultat. Une sous-naine O est membre d'une binaire spectroscopique dont l'analyse des vitesses radiales a pu être faite, mais la masse trouvée (4 en unités solaires, pouvant à la rigueur être réduite jusqu'à 1,2) est plus grande que la limite citée plus haut. On a également déterminé la masse de quelques sous-naines O à partir des paramètres de l'atmosphère et de la magnitude absolue bolométrique déduite de la distance, mais les incertitudes sont si grandes que la fourchette dans laquelle se situent les masses possibles est considérable. Dans un cas seulement (pour HD 49798) l'analyse a pu être nettement plus précise et la masse trouvée, 1.75 ± 1 en unités solaires, est un peu plus grande que les valeurs généralement admises.

La seule détermination directe du rayon d'une sous-naine O qui ait été faite dans une binaire à éclipse (BD −3° 5357, SdO + G8III) a donné $R \sim 0.1 \ R_\odot$. Par ailleurs, une évaluation du rayon de HD 49798 a donné $R = 1.45 \pm 0.25 \ R_\odot$. Enfin, les rayons impliqués par les masses et les gravités généralement admises pour les sous-naines O sont données dans la table I-31.

En ce qui concerne les étoiles de Wolf-Rayet, beaucoup se trouvent dans des systèmes binaires mais des difficultés considérables apparaissent dans le calcul des masses car les différentes raies donnent souvent des courbes de vitesses radiales différentes. La table I-32 donne les résultats obtenus à l'aide de 15 systèmes binaires avec ou sans éclipse, mais où les deux systèmes de raies sont mesurables. Les masses de la composante Wolf-Rayet sont données, la valeur de l'inclinaison i du plan de l'orbite ayant été estimée de différentes façons suivant les cas. La valeur moyenne des masses de la composante Wolf-Rayet est 17 en unités solaires; celle des rapports entre la masse de la composante Wolf-Rayet et celle de la composante O est 0.50.

Les déterminations directes des rayons des étoiles de Wolf-Rayet à partir des observations d'éclipses sont pratiquement impossibles à cause de la complexité de ces éclipses. Un effort considérable a cependant été fait pour interpréter les éclipses dans le système de V444 Cyg. Par ailleurs, le diamètre angulaire de la composante Wolf-Rayet de γ^2 Vel a pu être mesuré par des méthodes interférométriques, dans le continu à 4430 Å et dans la raie d'émission 4650 Å. Enfin, des méthodes indirectes analogues à celles utilisées pour les étoiles O normales peuvent être appliquées et donnent des rayons voisins de 10 en unités solaires, les sous-types WN7 et WN8 ayant cependant des rayons moyens supérieurs, environ 30 unités solaires.

Pour terminer, une liste de catalogues imprimés est donnée.

Deuxième Partie

Les étoiles de type O sont celles qui correspondent aux étoiles des types spectraux les moins avancés, aux étoiles les plus chaudes; les objets Of montrent en émission la raie de HeII 4686 et quelques fois d'autres raies d'émission, en plus du spectre normal d'absorption qui sert de base à la classification des étoiles de type O. Les étoiles de Wolf-Rayet (W-R), qui furent identifiées par Wolf et Rayet à l'Observatoire de Paris dès l'année 1867, sont également des objets chauds et lumineux avec des raies d'émission fortes et larges ($\sim 10^3$ km sec^{-1}) dans la région optique. Ces simples faits d'observation sont acceptés par les deux éditeurs de cette monographie, mais l'interprétation de ces détails et d'autres caractéristiques observées de ces objets, diffèrent largement d'un auteur à l'autre, comme nous le verrons.

Les auteurs de la partie II de cet ouvrage adoptent le point de vue selon lequel les spectres d'émission anormaux des étoiles Of et de Wolf-Rayet sont dus à un vent stellaire plus substantiel que celui qui caractérise les étoiles O à raies d'absorption. La force du vent elle-même est presque certainement associée à l'extrême température superficielle et à l'extrême luminosité de tous ces objets, mais d'autres paramètres physiques comme les différentes compositions chimiques, et les composantes non-thermiques de la production d'énergie, peuvent jouer un rôle. Les étoiles Of sont parmi les étoiles O les plus évoluées et les plus lumineuses. On pense que le phénomène Wolf-Rayet se produit pendant les phases d'évolution correspondant à la combustion de l'hélium, pour toutes les étoiles initialement plus massives qu'une valeur limite de l'ordre de $\sim 40 \ M_\odot$. Leur état évolutif peut être considéré comme représentatif d'une phase rouge et supergéante, non-observée par ailleurs, des étoiles les plus massives. (Le phénomène Wolf-Rayet est aussi présent dans quelques étoiles de masse beaucoup plus faible, "étoiles centrales" fortement évoluées des nébuleuses planétaires, objets qui ne seront pas discutés dans cette monographie.) Quelques centaines d'étoiles O et Of et d'objets de Wolf-Rayet sont connus dans notre propre Galaxie, et beaucoup d'autres ont été identifiés au sein des galaxies du groupe local, comme on le verra.

La partie II de ce volume discute de quatre thèmes associés les uns aux autres et relatifs à ces étoiles les plus chaudes et les plus lumineuses: le chapitre 2 donne une vue d'ensemble des observations relatives à la classification spectrale et aux propriétés extrinsèques des étoiles O et de Wolf-Rayet. Dans le chapitre 3, les paramètres intrinsèques de luminosité, température effective, masse et composition des étoiles, et une discussion de leur variabilité sont présentés. Dans le chapitre 4, les propriétés du vent stellaire sont résumées. Enfin, dans le chapitre 5 les questions voisines, concernant l'effet du rayonnement stellaire et des vents sur l'environnement circumstellaire immédiat, sont abordées. Vers la fin de la partie II de l'ouvrage, une synthèse des relations évolutives entre ces différentes étoiles sera suggérée.

Les étoiles O sont classées dans le système MK à deux dimensions, principalement à partir des caractéristiques d'absorption de l'hydrogène et de l'hélium, du type O jusqu'au type O9,5 avec des divisions en demi-classes commençant à O5,5. Les classes d'étoiles Of se localisent dans la zone des étoiles O intrinsèquement les plus brillantes, jusqu'au type 08.5; les supergéantes de type O9 et O9.5 montrent souvent des raies d'émission, mais pas la raie de HeII λ 4686. La distinction entre les étoiles Of et les supergéantes O9 et O9.5 est cependant plus apparente que réelle, puisque l'équilibre d'ionisation, dans les types O les plus tardifs, supprime la possibilité d'observer l'hélium ionisé. Un très petit nombre, probablement seulement les plus lumineuses des supergéantes de type B, ont des raies de Balmer présentant des caractéristiques d'émission. Les spectres d'absorption des étoiles O et Of prend sa naissance dans la photosphère stellaire; il est couplé d'une façon étroite avec les paramètres: température effective et luminosité. De même, ces spectres seront liés aux quantités plus fondamentales que sont la masse et la composition chimique, par l'intermédiaire des modèles de structure stellaire. L'observation des systèmes binaires confirme cette description d'ensemble.

Par contraste, la caractéristique principale du spectre optique des étoiles de Wolf-Rayet est la dominance des raies d'émission. La classification spectrale est unidimensionnelle, et divisée en deux sous-types majeurs, les étoiles WN, dans le spectre desquelles les raies des ions d'hélium et d'azote sont observables, et les étoiles WC, dont le spectre contient des raies de carbone et d'oxygène, ainsi que les raies des ions de l'hélium. Un sous-type supplémentaire peu peuplé des étoiles WO, dont le spectre comporte de fortes raies de l'oxygène VI a également été identifié. Par analogie avec le système MK, les sous-types "d'excitation" et "d'ionisation" peuvent être distingués les uns des autres, grâce à la mesure du rapport des intensités de diverse raies des ions de ces éléments. Ces sous-types vont de WN2 à WN9, de WC4 à WC9 et d'"avancés" à "tardifs", respectivement.

Dans un petit nombre d'étoiles WN, l'hydrogène semble présent, mais généralement le spectre des étoiles de Wolf-Rayet est remarquable par l'absence de cet élément. Et les raies d'absorption ne sont généralement pas présentes dans le spectre des étoiles de Wolf-Rayet, sauf dans quelques cas associés à l'existence d'un système binaire ou d'un compagnon plus distant, ou encore, en tant que profils de type P Cygni dus au vent associé à certaines caractéristiques en émission. Un spectre d'absorption dans les raies de Balmer les plus élevées semble n'être une caractéristique intrinsèque que de seulement un très petit nombre d'étoiles WN. Il est quelque peu difficile de séparer clairement "les étoiles Of les plus extrêmes" des "étoiles WN les moins extrêmes".

Le spectre d'émission des étoiles de Wolf-Rayet est caractéristique du vent stellaire, dont les propriétés sont couplées de façon seulement très lâche aux paramètres température effective et luminosité; en fait la relation entre les types spectraux des étoiles de Wolf-Rayet et les propriétés superficielles n'est pas bien établie. On peut supposer que ces étoiles W-N et WC, dans les sous-classes de la plus haute excitation (ou ionisation) auraient les températures effectives les plus élevées; mais ceci n'a pas encore été démontré de façon quantitative. Encore plus éloignée d'une compréhension simple est la constatation d'une relation entre les sous-classes et les masses stellaires, bien que les compositions, bien déterminées, des étoiles WN et WC, conduisent de façon naturelle à identifier ces étoiles comme résultant des combustions, respectivement, de l'hydrogène et de l'hélium. On s'attendrait ainsi à ce que les étoiles WC soient plus nettement évoluées que les sous-types WN, mais la relation des sous-classes d'excitation aux conditions initiales de masse et de composition, n'est pas claire. Il n'est pas encore possible de construire des modèles de structure stellaire capables de prédire les températures effectives ou les rayons des étoiles de Wolf-Rayet, à cause de l'absence d'une compréhension complète de la physique du vent.

Les étoiles ont été classées, historiquement, au moyen de leur spectre optique, puisque nos yeux et les détecteurs photographiques sont dans ce domaine les plus sensitifs des récepteurs, et que l'atmosphère terrestre est plus ou moins transparente à ces longueurs d'onde. Les progrès récents de la technique, et l'entrée en scène des détecteurs embarqués, nous ont maintenant donné des informations sur ce qui se passe aux autres longueurs d'onde. Aussi bien les longueurs d'onde de l'ultraviolet que celles de l'infrarouge ont été étudiées pour un certain nombre de ces étoiles, dont les spectres ont été analysés. Le spectre d'absorption des étoiles O se prolonge, en quelque sorte, vers d'autres longueurs d'onde, bien que s'y ajoutent, dans les régions de l'ultraviolet, les caractéristiques de type P Cygni liées au vent stellaire. Les étoiles de Wolf-Rayet retiennent leur spectre, principalement d'émission, dans les domaines ultraviolet et infrarouge.

Un petit nombre d'étoiles de type O se signale par des intensités anormales des raies des ions d'azote ou de carbone (type ON et OC). Les sous-types ON peuvent être associés à des différences de composition dues à leur structure, peut-être liées aux effets de mélange,—alors que les classes OC peuvent représenter des conditions extrêmes d'ionisation; mais ni l'un ni l'autre des deux groupes n'est encore bien compris. En dehors de ces quelques objets, les autres étoiles O étudiées jusqu'à présent dans notre Galaxie ont des compositions normales. Dans un travail récent de mes collègues Abbott, Bohannan, Hummer et Voels, il semble que les étoiles Of et les étoiles supergéantes O tardives et très lumineuses présentant des raies d'émission telles que Alpha Cam, peuvent avoir une abondance anormalement grande d'hélium. Ce matériau aurait alors été mélangé, à partir des régions du coeur où a lieu la combustion d'hydrogène, et les étoiles seraient alors quelque peu évoluées. Dans ce qui suit, je prendrai en considération les étoiles Of comme des étoiles O, pour en discuter les propriétés, sauf lorsque leur état avancé d'évolution et la présence de raies d'émission aura besoin d'être soulignée.

Les étoiles O et de Wolf-Rayet sont des objets que l'on peut cataloguer comme de "population I extrême", c'est-à-dire qu'ils sont pour la plus grande part confinés de façon étroite dans le plan galactique, et qu'ils définissent la structure des bras spiraux. la distribution en longitude montre en particulier une liaison étroite entre les étoiles de Wolf-Rayet et les étoiles les plus massives de type O. Les concentrations importantes de poussière dans le plan galactique nous empêchent d'identifier des étoiles de ce type à des distances sensiblement plus grandes que quelques kpc, mais ceci est encore beaucoup plus loin que la plupart des autres étoiles. Dans cette limite, nous avons une information raisonnablement complète sur les types et la fréquence des différents sous-types d'étoiles O et Wolf-Rayet.

De même, beaucoup d'étoiles O et de Wolf-Rayet ont été identifiées dans les galaxies extérieures, par exemple dans les Nuages de Magellan, dans M31, M33, NGC6822 et IC1613. Dans ces exemples de galaxies du groupe local, nous commençons à avoir les premières indications de différences possibles dans la fréquence et la distribution des étoiles de l'extrémité la plus haute de la série principale. Assez curieusement, les divers sous-types d'étoiles de Wolf-Rayet ne sont pas aussi bien représentés dans chacun de ces environnements: ainsi les étoiles WC tardives et les étoiles de type WC, ont-elles seulement été trouvées, dans le voisinage solaire, dans la direction du centre galactique. Cependant, les autres sous-classes sont trouvées, pas seulement dans le voisinage solaire, mais aussi bien dans d'autres galaxies. Finalement une telle information sur la distribution des étoiles massives nous aidera à comprendre la formation et l'évolution des galaxies, mais nous en sommes maintenant seulement à un niveau très primitif de notre compréhension. Le spectre des étoiles O dans le Petit Nuage de Magellan est clairement anormal, en ce qu'elles ont des raies plus faibles, cohérentes avec ce que l'on connaît actuellement sur la composition du gaz interstellaire; les étoiles O du Grand Nuage de Magellan ressemblent beaucoup plus à leurs équivalents galactiques.

Quelque 40% de la totalité des étoiles O sont des binaires, comme cela est mis en évidence par leur vitesse radiale périodique. La même fraction semble aussi s'appliquer aux étoiles Wolf-Rayet, ce qui suggère qu'aucun compagnon ne joue de rôle essentiel dans leurs anomalies, en tout cas dans la plupart des cas.

Les magnitudes absolues visuelles des étoiles O et de Wolf-Rayet peuvent être déterminées pour les objets identifiés dans les amas galactiques et dans les associations de distances "connues", comme dans les Nuages de Magellan. Les "calibrations" des magnitudes absolues M_v en type spectral sont décrites dans le chapitre 3, mais la détermination des luminosités correspondant à ces nombres est empoisonnée par l'incertitude sur l'échelle des températures effectives et des corrections bolométriques pour les étoiles de Wolf-Rayet. Les "meilleures" estimations de ces quantités sont données dans d'autres tables de ce chapitre, ainsi qu'une discussion des problèmes rencontrés dans la détermination de ces quantités, et des incertitudes associées.

Le Docteur Dietrich Baade a donné, dans la dernière partie du chapitre 3, une vue d'ensemble des variations intrinsèques des caractéristiques photosphériques. Le flux et la distribution du flux issus du spectre continu, et les caractéristiques d'absorption et d'émission sont sans doute variables à quelque niveau (pas toujours encore décelable) dans un grand nombre des étoiles étudiées (et peut-être dans toutes). Quelques-unes des variations les plus remarquables se rencontrent dans les systèmes de binaires serrées, dans lesquelles des interactions compliquées entre les étoiles et leur vent prennent place. Les supergéantes bleues les plus lumineuses (LBV) montrent souvent des variations facilement identifiables, potentiellement dues au fait que leur surface a des propriétés proches de la limite d'Eddington. Dans quelques-unes des ces supergéantes, des "sursauts" ou des évènements individuels violents ont été observés à quelques occasions, dans le passé.

Parmi les étoiles simples les plus proches de la série principale, la variabilité de basse amplitude peut être attribuée à des pulsations non-radiales (NRP) dans lesquelles divers modes peuvent être présents, impliquant le phénomène de "changement brutal de mode". Comme Baade le remarque, le travail d'observation systématique commence juste, et dans de nombreux cas, les effets sont très faibles et souvent juste à la limite de détectabilité. Des détecteurs de haute qualité et une amélioration du rapport signal-sur-bruit aideront d'une façon remarquable notre compréhension de la physique. L'étude de Baade concerne principalement le spectre photosphérique; la variabilité qu'il prend en considération se manifeste-t-elle directement dans les vents stellaires de ces objets? Peut-être, mais la relation entre ces deux phénomènes n'est pas encore très claire.

Le chapitre 4 contient plusieurs études individuelles de vents stellaires dans les étoiles O et de Wolf-Rayet. Katy Garmany y présente une étude détaillée de l'ensemble des paramètres des vents stellaires des étoiles de type O, en accordant une attention particulière au taux de perte de masse déduits des mesures, et aux vitesses terminales, qui peuvent être obtenues à partir des données de la façon la plus directe. La quantité dont le taux de perte de masse dépend de la façon la plus importante est la luminosité stellaire; bien que d'autres propriétés stellaires telles que la masse, la gravité, la température effective puissent jouer un rôle, ceci n'a pas encore été montré de façon quantitative. La relation sans ambiguïté entre le taux de perte de masse et la luminosité, suggère fortement que la pression de radiation est le facteur déterminant des propriétés du vent stellaire des étoiles O (et des supergéantes de type B).

La relation entre le taux de perte de masse et la luminosité dans les étoiles de Wolf-Rayet est moins claire, bien qu'ici l'incertitude dans la détermination de la correction bolométrique puisse affecter la description des phénomènes. Les étoiles de Wolf-Rayet ont des vents beaucoup plus forts que les étoiles O avec des taux de perte de masse de quelques 10^5 masses solaires par an. Les types les plus avancés ont des vitesses terminales quelques peu plus élevées que les types les plus tardifs, comme ceci peut aussi être déduit de la largeur des raies d'émission optiques, mais les valeurs trouvées ne sont pas si grandes que celles que l'on trouve dans le cas des étoiles O. La densité des vents est plus forte dans les étoiles de Wolf-Rayet, ce qui conduit à ce que la profondeur optique unité corresponde à des régions où la matière est déjà en mouvement. Ceci conduit directement à la compréhension du fait que le spectre d'émission dominant résulte directement de l'importante densité du vent.

R. P. Kudritzki, A. Pauldrach et J. Puls donnent une vue d'ensemble des aspects essentiels de la *théorie* des vents poussés par la radiation et comparent ces théories avec les données détaillées de l'observation. Cette théorie, telle qu'elle a été développée au cours des ans par un grand nombre d'astrophysiciens, a été couronnée de remarquables succès dans ses principales caractéristiques. Il y a une correspondance raisonnablement bonne pour les structures générales des vents stellaires, par exemple dans la prévision de la relation entre ces vents et la luminosité, comme dans leur relation avec les vitesses terminales observées. Dans les travaux plus récents que ce qui est reporté dans cette monographie, la théorie a été utilisée par Kudritzki et ses collègues, et par Abbott, pour prévoir la façon dont le taux de perte de masse et la vitesse terminale dépendent de la composition chimique d'ensemble. Récemment des observations des étoiles O des Nuages de Magellan, rapportées par Garmany et Conti, ont confirmé les prévisions des vitesses terminales les plus basses dans le Petit Nuage de Magellan, mais on n'a pas trouvé un taux de perte de masse plus petit. Une explication possible pourrait être la détermination médiocre des luminosités des étoiles du Nuage de Magellan, liées aux incertitudes sur la relation entre le type spectral et la température effective. Ce débat important reste actuellement ouvert.

Une exception très notable à la théorie généralement formulée des vents stellaires entraînés par le rayonnement est leur variabilité observée. Ceci peut seulement se comprendre en termes d'une théorie d'un vent poussé par la radiation *si* la radiation stellaire elle-même varie, ou *si* des instabilités se produisent naturellement dans le vent. Il serait important de démontrer que le genre de détails subtils dans la variabilité, tels que ceux trouvés dans les photosphères stellaires et décrits par Baade dans le chapitre 3, se reflètent d'une façon directe dans les vents stellaires. Jusqu'à présent, il n'y a pas eu de travail quantitatif sur cette question pour les étoiles O et Wolf-Rayet, mais ceci serait une voie prometteuse à explorer. Le goulot d'étranglement est probablement du côté de l'observation, et il est nécessaire de s'attaquer d'une façon obstinée et concentrée à l'étude d'un certain nombre d'étoiles particulièrement typiques, dans les différents domaines de longueur d'onde.

La variabilité du vent stellaire, particulièrement dans les composantes "étroites" fréquemment rencontrées dans l'absorption des profils P Cygni, est décrite par Huib Henrichs dans le chapitre 4. Après sa discussion des données, tout particulièrement des échelles de temps, il considère les différents modèles proposés pour expliquer l'existence de ces caractéristiques étroites. La plupart d'entre eux ne peuvent expliquer les données d'aucune façon satisfaisante. L'auteur est en faveur d'une liaison avec le phénomène NRP, en particulier le changement brutal de mode, comme une façon d'injecter de l'énergie non-thermique dans le vent poussé par la radiation. Bien que beaucoup plus de travail, aussi bien dans le domaine de l'observation que dans celui de la théorie, devra être consacré à cette question, mon impression personnelle est que ceci apparaîtra comme la façon correcte d'aborder le problème. Ainsi les vents des étoiles O (et de Wolf-Rayet ?) sont-ils essentiellement poussés par le rayonnement, mais une énergie supplémentaire modifiant la structure d'ensemble de ces vents peut venir de phénomènes superficiels tels que les pulsations non-radiales.

Une étoile est une entité gazeuse qui se maintient grâce à un équilibre subtil entre la force de gravité (vers le bas) et la force qu'exerce (vers l'extérieur) la pression de radiation liée aux combustions nucléaires au coeur de l'étoile. La "surface" (ou photosphère) est la région dont le rayonnement sort. Mais l'étoile ne "finit" pas à ce point, puisqu'il existe encore de la matière à l'extérieur. Dans les étoiles de cette monographie, une quantité *substantielle* de matière se trouve au-dessus de la photosphère, sous la forme de vent stellaire. Ce vent est en interaction avec la matière localisée au-delà de l'étoile, soit du fait de l'histoire totale antérieure de la perte de masse, soit par l'intermédiaire d'"évènements" épisodiques possibles, ou encore à partir de la matière qui se trouvait initialement autour de l'étoile. Cette matière issue partiellement d'éjections par l'étoile, et partiellement de l'accrétion de matériau interstellaire, a été modifiée de façon substantielle par l'étoile elle-même, que ce soit par son rayonnement intense ou par le moment transporté par son vent. Je me suis référé à cette matière comme au milieu stellaire *local*, ou LSM, afin de le distinguer du matériau interstellaire relativement non modifié.

Les étoiles O et Wolf-Rayet affectent notablement leur environnement; pendant le processus de formation d'étoiles, l'apparition d'une ou d'un plus grand nombre d'étoiles O récemment nées affecte de façon dramatique les naissances ultérieures d'étoiles. Leur disparition (comme supernova) affectera aussi de façon considérable l'environnement et altèrera de façon inexorable ce qui arrivera au voisinage. En plus de ces évènements exceptionnels, l'étude de l'interaction de l'étoile avec son environnement pendant sa vie de quelques millions d'années, est l'un des plus intéressants des phénomènes qui ont émergé récemment de l'étude des étoiles chaudes. Un grand nombre de processus physiques sont très loin d'être linéaires et vraisemblablement également non-thermiques. La plupart du travail fait jusqu'à présent est parti de l'hypothèse de la symétrie sphérique et d'un milieu homogène. Il est clair que le milieu interstellaire est hautement non-homogène; le milieu stellaire local devrait avoir ce caractère de façon encore plus nette puisqu'il a été modifié par le vent stellaire et par le champ de rayonnement. Des exemples de cette nature sont les "anneaux" et les "bulles" associés à un petit nombre d'étoiles O et Wolf-Rayet. Les conditions au voisinage d'autres étoiles ne peuvent pas être propices à la détection facile des ces structures.

L'un des principaux objectifs du travail décrit dans la partie II de cette monographie, est la confirmation de l'hypothèse que les étoiles de Wolf-Rayet sont les descendants hautement évolués des étoiles les plus massives. Aussi bien les observations que les modèles théoriques conduisent à cette conclusion essentielle. Des arguments indépendants viennent de différentes voies d'exploration.

Il est bien connu que les étoiles massives perdent de la masse, étant donné la présence évidente de profils P Cygni dans les raies de résonance ultra-violettes; de plus, autour de nombreuses étoiles Of, et de quelques variables bleues lumineuses (LBV), on trouve un matériau enrichi en azote, qui a dû avoir été éjecté de l'étoile. Il est raisonnable de conclure que la perte de masse et les brassages de matière stellaire doivent jouer un rôle dans l'évolution des étoiles massives puisque celles-ci contiennent encore de l'hydrogène et qu'elles sont encore en train de brûler cet hydrogène en leur intérieur.

Les étoiles supergéantes rouges sont connues pour être des objets dont le coeur brûle de l'hélium; cependant on n'en 'a pas encore trouvé de plus brillantes qu'une magnitude bolométrique d'environ -10, ce qui correspond à des étoiles dont la masse initiale serait voisine de 40 $M_\odot$. Cependant des supergéantes bleues, bolométriquement plus brillantes, sont connues, et l'on a trouvé des étoiles plus massives; celles-ci doivent poursuivre leur phase d'évolution par combustion de l'hélium d'une façon différente qu'une supergéante rouge.

Les étoiles de Wolf-Rayet sont observées comme ayant les propriétés de composition anormale en hélium et en azote, ou en hélium carbone et oxygène, et doivent contenir très peu, ou dans la plupart des cas pas du tout d' hydrogène. Celles des étoiles qui sont membres de systèmes binaires sont observées comme plus lumineuses et pourtant moins massives que leurs compagnons. Une affirmation selon laquelle les étoiles de Wolf-Rayet sont les descendants, dans leur phase de combustion de l'hélium, des étoiles les plus massives, s'accorde ainsi agréablement avec tous les faits observés donnés ci-dessus.

Une confirmation indépendante de cette description peut être trouvée dans la concentration galactique générale des étoiles O les plus massives, des supergéantes B *et* des étoiles de Wolf-Rayet, dans les bras spiraux dirigés vers le centre galactique; les étoiles moins massives de type O ne présentent pas une telle distribution en longitude, et les étoiles B de la série principale ou les géantes B, sont même encore moins comparables aux étoiles O les plus massives, en ce qui concerne leur distribution.

Des modèles d'étoiles prenant en compte la perte de masse et le mélange convectif sont capables de produire des noyaux complètement dépouillés dans les phases de combustion d'hélium. Ces objets se trouvent dans la partie relativement chaude du diagramme HR, plutôt que dans les portions de ce diagramme où se trouvent les supergéantes rouges plus froides. L'apparence énigmatique des spectres, purement d'émission, peut être reproduite à partir de forts vents stellaires, résultant eux-mêmes de valeurs extrêmes de la luminosité et de la température, valeurs que l'on s'attendrait à trouver pour des objets massifs, en train de brûler leur hélium, et où peu d'hydrogène resterait, voire éventuellement pas d'hydrogène du tout.

L'interprétation des étoiles de Wolf-Rayet en tant qu'objets brûlant leur hélium est en accord avec toutes les données de l'observation, et reste cohérente avec les prévisions de la théorie de l'évolution stellaire.

Les étoiles Of semblent être les phases ultimes de l'évolution des étoiles massives, dans leur période de combustion de l'hydrogène; les modèles stellaires suggèrent que ces étoiles peuvent déjà être en train de tourner "vers la gauche" dans la partie élevée du diagramme HR. La masse peut également avoir été perdue dans des évènements épisodiques associés aux étoiles LBV, qui jouent peut-être un rôle intermédiaire entre les phases de combustion d'hydrogène et de l'hélium dans certaines étoiles massives (peut-être dans toutes?). Cette connexion n'est cependant pas encore certaine. Les binaires jouent probablement seulement un rôle mineur dans l'évolution des étoiles massives, à moins que les compagnons ne soient initialement très proches l'un de l'autre, puisque les modèles indiquent que les étoiles vont tourner dans le diagramme HR vers la "gauche", alors que le coeur brûle encore de l'hydrogène, et avant que le rayon stellaire ne devienne important.

Les types et la distribution des différentes sous-catégories d'étoiles de Wolf-Rayet varient d'un endroit à l'autre, mais ces relations ne sont pas encore bien établies en raison de la difficulté qu'il y a à prédire, à partir des modèles stellaires, les températures effectives. Il se peut que ces étoiles aient la possibilité de nous donner une certaine information en ce qui concerne les détails de l'évolution des étoiles les plus massives dans des environnements variés.

Epilogue. Les contributions de la partie II de cette monographie ont été essentiellement terminées à la fin de 1986. Au début de 1987, le rayonnement issu de l'explosion de la supernova du grand Nuage de Magellan a atteint la Terre, nous donnant une vue sans précédent de la mort violente d'une étoile massive. L'objet progéniteur immédiat, NS –69 202, était une étoile B3 supergéante; il existe une évidence circonstantielle pour une histoire antérieure de cet objet en tant que supergéante rouge. La masse initiale de l'objet, déduite de la luminosité bolométrique de l'explosion, était d'environ 20 $M_\odot$, correspondant à un type O9 sur la série principale.

Notre connaissance antérieure de l'évolution stellaire aurait suggéré qu'une telle étoile de 20 $M_\odot$ terminerait sa vie en tant que supergéante rouge. Mais pour des quelque raison pas encore complètement explicitée, l'objet NS –69 202 est retourné, avant l'explosion, sur la région bleue du diagramme HR. En tant que supergéante bleue, cet objet a un rayon relativement petit; ainsi le rayonnement optique était-il relativement faible par rapport à ce qu'il fut pour d'autres supernovae de type II. A posteriori, il a été possible de paramétriser la perte de masse et le mélange convectif dans une série de modèles en évolution de 20 $M_\odot$, dans lesquels ce scénario pourrait se produire normalement dans des conditions initiales similaires. On prédirait alors que d'autres supernovae relativement faibles de type II, comparables à 1987a, devraient être des objets communs. Jusqu'à présent, il n'en a pas été identifié de tels dans d'autres galaxies, mais il n'est pas exclu qu'un effet de sélection ait empêché leur détection.

L'onde de choc associée au phénomène de supernova est actuellement en train de se propager vers l'extérieur, à travers le milieu stellaire local de l'étoile explosée. En même temps qu'elle avance, elle rencontre, et excite, de la matière éjectée antérieurement par les phases supergéante bleue et supergéante rouge qui pourraient avoir été distinguées l'une de l'autre par leur vitesse "terminale" différente. Les observations nous révèleront donc, avec le temps, et dans une certaine mesure, une portion de l'histoire passée de NS -69 202, et en particulier sa composition. Il ne serait pas très surprenant de trouver un accroissement de l'azote dans ces stades évolutifs; la détection d'une abondance anormale d'hélium sera probablement plus difficile du fait des conditions physiques du nuage gazeux, au moment où le choc l'atteint. La supernova 1987a sera certainement un objet révélateur, au cours des années à venir, au fur et à mesure de la mise en évidence des conditions physiques de plus en plus anciennes de son histoire. Alors que la matière est en expansion à partir de la supernova, elle devient moins opaque, révélant le coeur de l'étoile tel qu'il était dans une époque précédant immédiatement l'explosion. Une information nouvelle, et probablement sans précédent, nous attend certainement dans le dévoilement progressif de la physique d'une étoile massive. Notre compréhension de telles étoiles massives pourrait bien être complètement modifiée par ces observations et par leur interprétation.

Troisieme Partie

Dans la troisième partie, on montre que l'on ne peut expliquer l'intensité des raies d'émission ni celle de la grande majorité des raies d'absorption fortes du spectre des étoiles O et Wolf-Rayet si l'on utilise uniquement la théorie classique des atmosphères stellaires en équilibre hydrostatique, radiatif et statistique; il en est de même pour l'émission de rayons X, les excès de rayonnement infrarouge et radio, l'existence d'un vent stellaire. Le problème général de l'interprétation du spectre des étoiles O, Of et Wolf-Rayet est posé dans le chapitre 6. Dans le chapitre 7, on donne un exposé de la théorie des atmosphères des étoiles chaudes telle qu'elle résulte des publications faites jusqu'à la fin de 1983, avec ses succès et ses échecs. Dans le chapitre 8 on montre comment le fait d'ajouter à l'ensemble des hypothèses de base de la théorie des spectres stellaires celle de l'existence de champs magnétiques inhomogènes relativement faibles peut aider beaucoup à comprendre les observations.

Des mouvements de matière produits par la pression de radiation peuvent expliquer quelques uns des caractères spectroscopiques observés, mais il semble très probable que les hautes températures et les flux de matière observés dans le manteau des étoiles O, Of et Wolf-Rayet soient dus principalement au chauffage et aux transferts de quantité de mouvement qui se produisent lorsque des zônes de turbulence dans la photosphère réagissent sur les points de départ de lignes de force magnétiques traversant celle-ci. Dans le chapitre 8 on admet que les lignes de force magnétiques nécessaires sont créées par des effets dynamo ayant leur origine dans la couche convective He^+ située dans les parties les plus profondes de la photosphère des étoiles O et Wolf-Rayet. Les mouvements de turbulence invoqués pour expliquer la production de ces ondes magnétohydrodynamiques (MHD) qui se propagent ensuite à travers le manteau des étoiles O, Of et Wolf-Rayet, sont dus en définitive à la rotation de l'étoile. On admet que la viscosité du fluide en rotation qui forme la surface de l'étoile produit un mouvement helicoïdal turbulent dans la photosphère parallèlement à la surface de l'étoile. C'est ce flux turbulent qui serait à l'origine des effets dynamo de la zone convective He^+ créant les ondes MHD.

Il peut se produire des mouvements de turbulence lorsqu'un flux de radiation traverse un gaz ionisé à faible densité dans lequel existe un gradient de vitesse (voir chapitre 7). Mais il reste à démontrer que les ondes accoustiques nécessaires dans la théorie de ce phénomène puissent se propager depuis la photosphère ou depuis les régions plus profondes où elles peuvent se former jusque dans le manteau d'une étoile. Les gradients de vitesse nécessaires peuvent être produits par l'action des ondes MHD (voir chapitre 8). À partir du moment où un gradient de vitesse suffisant a été créé, les effets de la pression de radiation dans les raies peuvent accroître le gradient de vitesse créé par les ondes MHD.

Cinq questions sont posées à la fin du chapitre 6. De brèves réponses sont données dans ce qui suit:

1. La source de l'énergie non radiative et de la quantité de mouvement qui produisent la formation du manteau est la rotation de l'étoile.

2. L'énergie et la quantité de mouvement provenant de la rotation de l'étoile sont transmis au gaz du manteau par des ondes MHD se propageant le long de lignes de force magnétiques qui traversent le manteau. Ces lignes de force magnétiques sont le résultat des effets dynamo engendrés par les mouvements hélicoïdaux dus à la rotation de l'étoile, en présence du champ magnétique faible supposé présent dans la photosphère. Il est probable que les lignes de force du champ magnétique ne traversent la photosphère que dans des régions étroites concentrées aux bords des éléments de turbulence, comme dans le cas du Soleil.

3. La libération d'énergie non radiative et de quantité de mouvement due aux ondes MHD est peu importante dans la photosphère car dans les parties de la photosphère où se trouvent les lignes de force du champ magnétique, le rapport entre la pression gazeuse et la pression magnétique est grand. Dans les parties de la photosphère sans lignes de forces magnétiques, il n'y a pas d'ondes MHD.

4. Les phénomènes non radiatifs dans les atmosphères stellaires sont importants pour comprendre l'évolution des étoiles chaque fois que les critères empiriques choisis pour définir le type spectral sont influencés par la présence d'énergie et de quantité de mouvement d'origine non radiative dans le manteau. On admet souvent que le type spectral peut être utilisé comme un indicateur non ambigu des paramètres fondamentaux de la théorie de l'évolution stellaire (masse, rayon, température effective et composition chimique de l'étoile). Dans le chapitre 7 on montre que ces paramètres fondamentaux ne peuvent pas être déterminés sans ambiguïté à partir des spectres des étoiles O, Of et de Wolf-Rayet en utilisant uniquement la théorie classique étant donné que l'apparence des spectres dépend notablement de la quantité d'énergie non radiative apparaissant dans le manteau de l'étoile, de l'importance de ce manteau et des mouvements qui s'y produisent. Il est donc essentiel de prendre en compte tous les phénomènes non radiatifs si l'on veut décrire correctement l'évolution des étoiles très chaudes. Des renseignements sur les masses, les rayons et les températures effectives des étoiles O et Wolf-Rayet se trouvent dans la partie I de ce livre. Les études théoriques discutées dans la partie III indiquent que les étoiles O, Of et de Wolf-Rayet de Population I ont une composition chimique normale de type solaire. Dans quelques étoiles O sous-lumineuses, l'abondance relative de l'hélium par rapport à l'hydrogène est anormale. Il semble que dans certains cas une séparation d'origine gravitationnelle s'est produite alors que dans d'autres cas l'enveloppe riche en hydrogène a été perdue.

5. L'observation des phénomènes non radiatifs dans les atmosphères des étoiles O, Of et de Wolf-Rayet ne donne pas de renseignements sur le stade d'évolution individuel de ces étoiles.

Des spectres classés O, Of et Wolf-Rayet se rencontrent parmi les étoiles jeunes massives de Population I et parmi les étoiles sous-lumineuses que l'on a de bonnes raisons de faire correspondre à des stades très avancés de l'évolution stellaire. Les informations données dans la partie III de ce livre indiquent que les phénomènes spectroscopiques qui caractérisent les spectres O, Of et Wolf-Rayet sont dus très probablement au chauffage et au transfert de quantité de mouvement produits dans le manteau par la rotation d'une étoile chaude ayant un champ magnétique faible dans ses couches extérieures. Ce champ magnétique peut être primordial ou avoir été créé au cours de l'évolution de l'étoile. Tant que la théorie de l'évolution stellaire n'aura pas atteint un degré de précision suffisant pour que l'on puisse prédire l'apparition d'un champ magnétique dans les couches externes d'une étoile chaude, et plus précisément dans la zône convective He^+ de sa photosphère, il sera impossible d'utiliser des observations spectroscopiques qui sont liées au chauffage et au transfert de quantité de mouvement produits par l'action des ondes MHD dans le manteau pour prédire le stade d'évolution de l'étoile. Dans l'état actuel de la théorie, des informations supplémentaires, telles que la composition de l'atmosphère de l'étoile, sa luminosité, sa température effective et sa masse, sont nécessaires pour pouvoir estimer le stade d'évolution.

En résumé, l'apparence des spectres des étoiles O, Of et de Wolf-Rayet est fortement influencée par les phénomènes de chauffage et de transfert de quantité de mouvement non radiatifs se produisant dans le manteau. Il en résulte que le type spectral O, Of ou Wolf-Rayet ne suffit pas dans l'état actuel de la théorie pour connaître le stade d'évolution de l'étoile.

Il est possible que l'on puisse obtenir des informations intéressantes sur le stade de l'évolution à partir des caractères spectraux O, Of et Wolf-Rayet lorsque la théorie de l'évolution stellaire aura été étendue de telle façon que l'on puisse prévoir, d'une part l'existence de champs magnétiques dans la zône convective He^+ pour les étoiles chaudes massives (et également pour celles de faible masse), d'autre part les phases pendant lesquelles la rotation de l'étoile est suffisamment rapide pour que des mouvements tourbillonnaires se produisent dans les couches externes. Ces mouvements tourbillonnaires en présence des champs magnétiques sub-photosphériques créeraient des effets dynamo ayant pour conséquence l'expulsion de lignes de forces magnétiques hors de l'étoile et la formation d'un manteau dont les propriétés seraient ainsi prévisibles. Dans le chapitre 8 on explique comment l'expulsion de lignes de force magnétiques peut rendre compte du manteau des étoiles O, Of et de Wolf-Rayet, et également du taux de perte de masse.

SUMMARY

Part I

In Part I of this book, prepared before 1984, basic information is given about O and Wolf-Rayet stars indicating how these stars are defined and what their chief observable properties are. The frontier between "observable" and other properties is clearly rather arbitrary. It is linked closely with the personal idea one forms about the value of the theories used to interpret any phenomenon. For a long time the wavelength of a spectral line has been considered to be an observable quantity; the effective temperature has also been so acknowledged, especially since flux measurements have been extended into the ultraviolet. On the other hand, we consider that the age of a star is a result of the theory of stellar evolution. The relationship between the theory of stellar evolution and the observations is not discussed in Part I of this book.

O-type stars are defined by the character of their spectra in the visible range. A very blue distribution of energy and the presence of absorption lines due to ionized helium are defining characteristics. However, the O stars do not form a homogeneous group and stars having similar spectra can have masses, radii, ages, and internal structures which differ greatly. Most O stars known at this time are luminous, massive stars concentrated near the galactic plane. They are in the first stages of stellar evolution and are called "normal O stars." However, more and more O stars which are subluminous and have small masses are being discovered in the galactic halo and are called the "subdwarf O stars." Their internal structure is still poorly understood but it is certain that these hot, low-mass stars correspond to advanced stages of stellar evolution. The Wolf-Rayet stars are recognized by their spectra which are dominated by strong, wide emission lines due to highly ionized atoms.

Given these definitions, we describe the spectral classification, intrinsic colors, effective temperatures, masses, and radii of O and Wolf-Rayet stars as well as the principal studies leading to these results. So far as each observable quantity is concerned, the normal O stars, the subdwarf O stars, and the Wolf-Rayet stars present entirely different problems. Therefore, these three types of hot stars are treated separately.

Part II

The O-type stars are the hottest (earliest) spectral types; Of objects have HeII 4686 and sometimes other emission lines present, in addition to the normal absorption line spectrum upon which the O-type classification is based. Wolf-Rayet stars, first identified by Wolf and Rayet at the Paris Observatory in 1867, are also hot luminous objects with strong, broad ($\sim 10^3$ km sec^{-1}) emission lines in the optical region. These simple observational facts are agreed on by the two editors of this monograph, but the interpretations of these details and others that follow are vastly different.

The Part II authors adopt the point of view that the anomalous emission line spectra of Of and Wolf-Rayet stars are due to a stellar wind that is more substantial than is found in the absorption line O stars. The wind strength itself is almost certainly related to the extreme surface temperature and luminosity of all of these stars but other physical parameters such as composition differences and nonthermal energy input may well play a role. The Of stars are among the most highly evolved and luminous O stars. It is believed that the Wolf-Rayet phenomena are found during the normal helium burning stages of evolution of all stars that are initially more massive than some limit near $\sim 40\ M_{\odot}$. Their evolutionary status can be thought of as representing an otherwise not-observed red supergiant phase of the most massive stars. (The Wolf-Rayet phenomenon is also present in some of the much lower mass, highly evolved "central stars" of planetary nebulae which will not be discussed here.) Some few hundred individual O and Of stars and Wolf-Rayet objects are known in our Galaxy and many others have been identified among galaxies of the Local Group.

Part II of this volume discusses four related themes pertaining to these hottest and most luminous stars. Chapter 2 will deal with an observational overview of the spectroscopic classification and extrinsic properties of O and Wolf-Rayet stars. In Chapter 3 the intrinsic parameters of luminosity, effective temperature, mass and composition of the stars, and a discussion of their variability, will be presented. In Chapter 4, the stellar wind properties will be summarized, and in Chapter 5, the related issues concerning the effects of the stellar radiation and winds on the immediate interstellar environment will be given. Near the end of Part II, a summary of the evolutionary connections among these stars will be offered.

The O stars are classified in the two-dimensional MK system mostly from the hydrogen and helium absorption line features, starting with type O3 and ending at O9.5, with the intermediate (O.5) subclasses beginning at O5.5. The Of classification is found in the intrinsically brighter O stars up to type O8.5; O9 and O9.5 supergiants often show emission lines but not $\lambda 4686$ of He II. The distinction between Of and supergiant O9, O9.5 stars is more apparent than real as the ionization balance in the later O types suppresses the appearance of ionized helium. (A very few, probably only the most luminous, B-type supergiants show emission in the Balmer series.) The absorption line spectra of the O and Of stars originates in the stellar photospheres and it is closely coupled to the parameters of effective temperature and luminosity. These parameters are likewise tied to the more fundamental quantities of mass and chemical composition through the use of stellar structure models. Observations of binary systems confirm the overall picture.

By contrast, the main characteristic of the optical spectrum of Wolf-Rayet stars is the dominance of emission lines. The spectral classification is one-dimensional and it is divided into two major subtypes, the WN, in which lines of helium and nitrogen ions are seen, and the WC, which contain lines of carbon and oxygen along with the helium ions. An additional sparsely populated subtype, WO, with strong OVI lines has also been identified. By analogy with the MK system, "excitation" or "ionization" subtypes can be segregated by the ratios of various line strengths of ions of these elements. These range from WN2 to WN9, and WC4 to WC9, and from "early" to "late," respectively.

In a few WN stars, hydrogen appears to be present, but generally the spectra of Wolf-Rayet stars are notable for the absence of this element. Absorption lines are generally not found in Wolf-Rayet stars, except in some instances due to a binary or a more distant companion, or as P Cygni profiles of the wind associated with certain emission features. An upper Balmer line absorption spectrum appears to be an intrinsic feature in only a very few WN stars. There is some difficulty in clearly delineating "the most extreme Of stars" from the "least extreme WN stars."

The Wolf-Rayet emission line spectra are features of the stellar wind, whose properties are only loosely coupled to the parameters of effective temperature and luminosity; in fact, the relationship between the Wolf-Rayet spectral types and the surface properties is not well established. It is supposed that those WN and WC stars with the highest excitation (ionization) subclasses would have the highest effective temperatures, but this has not yet been demonstrated quantitatively. Even more poorly understood is the relationship of the subclasses and the stellar masses, although the well-determined compositions of the WN and WC spectra do lead in a natural way to identification of the stars as spawning products of core-hydrogen and core-helium burning material. Thus, one would expect that the WC stars are more highly evolved than the WN subtypes, but the relationship of the excitation subclasses to the initial conditions of mass and composition is not clear. Stellar structure models cannot be made that predict the effective temperatures or radii of Wolf-Rayet stars, because of a lack of complete understanding of the physics of the wind.

Historically, stars have been classified by their optical spectra, since our eyes and photographic detectors are most sensitive in this range and the Earth's atmosphere is more or less transparent at these wavelengths. Recent advances in technology, and the advent of detectors in space, have now given us new information at other wavelengths. Both the UV and IR wavelengths have been studied for a number of these stars and their spectra analyzed. The nominal absorption line spectrum of O stars is also found at other wavelengths, although with the addition of ubiquitous P Cygni features from the stellar wind in the UV regions. Wolf-Rayet stars retain their predominately emission-line spectrum in the UV and IR wavelengths.

A few O-type stars show anomalous line strengths of nitrogen or carbon ions (the ON and OC types). The ON subtypes may have self-induced composition differences, possibly due to mixing, and the OC classes may represent extreme ionization conditions, but neither group is as yet understood. Aside from these few objects, other O stars that were studied in our Galaxy have normal compositions. In recent work by investigators Abbott, Bohannan, Hummer, and Voels it appears the Of stars, and those very luminous late O supergiants with emission lines such as Alpha Cam, may have enhanced helium abundance. This material would have been mixed from the core-hydrogen burning regions and the stars would be somewhat evolved. In the following chapters, I will consider the Of stars as O types for discussion purposes, except where their advanced evolution stage and presence of emission lines need to be underlined.

O and Wolf-Rayet stars are "extreme Population I" objects (i.e., they are for the most part narrowly confined to the galactic plane and define its spiral arm pattern). The longitude distribution in particular shows an intimate connection between the Wolf-Rayet stars and the most massive of the O types. The heavy concentrations of dust in the galactic plane prevents us from identifying all such stars at distances substantially greater than a few Kpc, but this is further away than most other stars. Within these distances in our Galaxy, we have reasonably complete information on the types and numbers of the various O and Wolf-Rayet subtypes.

Similarly, many O and Wolf-Rayet stars have been identified in external galaxies (e.g., the Magellanic Clouds, M31, M33, NGC6822, and IC1613). In these examples of Local Group galaxies, we are beginning to see the first hints of possible differences in the numbers and distributions of stars at the upper end of the main sequence. Oddly enough, the various Wolf-Rayet subtypes are not equally represented in each of these environments: the late WC stars, the WCL types, have only been found in the solar vicinity inward towards the galactic center. Yet the other subclasses are found not only in the solar vicinity but in the other galaxies as well. Ultimately, such information on the distribution of massive stars will help us to understand the formation and evolution of galaxies, but we are now at only a primitive level of understanding. The spectra of the O stars in the SMC are clearly anomalous in having weaker lines, consistent with what is known concerning the composition of the interstellar gas: the LMC O stars are more like their galactic counterparts.

Some 40 percent of all O stars are binaries, as evidenced by their periodic radial velocities. Also, this fraction seems to be found among the Wolf-Rayet objects, suggesting that a companion star, in most cases, does not play a leading role in their anomalies.

Absolute visual magnitudes of O and Wolf-Rayet stars can be obtained for those objects identified in galactic clusters and associations with "known" distances and from members of the Magellanic Clouds. "Calibrations" of the M_v with spectral types are presented in Chapter 3, but the derivation of the luminosities from these numbers is highly approximate due to the uncertainty in the effective temperature scale and bolometric corrections for the Wolf-Rayet stars. "Best" estimates of these quantities are given in other tables in this chapter, along with a discussion of the problems and uncertainties in deriving the numbers.

Dr. Dietrich Baade has given an overview of the intrinsic variations of photospheric features in the last portion of Chapter 3. The flux and flux distribution from the continuous spectrum and the absorption and emission features are believed to vary at some level (not yet always detectable) in many (and maybe in all) stars. Some of the most noticeable variations are found in close binary systems in which complicated interactions between the stars and their winds are occurring. The most Luminous Blue supergiant stars (LBV) often show readily identifiable variations, potentially due to their surfaces having properties close to the Eddington Limit. In some of these supergiants, "outbursts" or singular violent events have been observed on occasion in the past.

Among single stars nearer the main sequence, low-level variability may be attributable to Nonradial Pulsations (NRP) in which various modes or the phenomena of "mode switching" may be present. As Baade notes, the systematic observational work is just beginning; as in many cases, the observational effects are subtle and often just at the limit of detectability. Advanced detectors and increased signal-to-noise data will dramatically help our understanding of the physics. Baade's review mostly concerns the photospheric spectrum of a star: Does the variability he considers manifest itself directly in the stellar winds of these objects? Possibly, but the connection is not yet clear.

Chapter 4 contains several individual reviews of the stellar winds of O and Wolf-Rayet stars. Dr. Katy Garmany gives a detailed review of the overall parameters of the stellar winds of O-type stars, paying particular attention to those inferred mass-loss rates and terminal velocities which can be most readily obtained from the data. The most important dependence of the mass-loss rate is upon the stellar luminosity; although other stellar properties such as mass, gravity, and effective temperature might play a role, this has not yet been shown quantitatively. The unequivocal relationship between mass-loss rate and luminosity strongly suggests that radiation pressure is the leading determinant of the stellar wind properties of O stars (and B supergiants).

The relationship between mass-loss rate and luminosity among the Wolf-Rayet stars is less clear because of the uncertainty in the bolometric correction. Wolf-Rayet stars have much stronger winds than O stars with mass-loss rates of some few 10^5 solar masses a year. The earlier types have somewhat larger terminal velocities than the later types, as can be inferred from their optical emission-line widths, but the numbers are no larger than what are found in O stars. The wind densities are higher in Wolf-Rayet stars, leading to optical depth unity being found where the material is already flowing. This leads directly to an understanding that the dominant emission-line spectrum is a direct result of the substantial wind density.

Drs. R.P. Kudritzki, A. Pauldrach and J. Puls give an overview of the essential features of the radiatively driven wind *theory* and a comparison to the observational details. This theory, as it has been developed over the years by a number of astrophysicists, has been remarkably successful in its major outlines. There is good correspondence in the overall features of stellar winds; e.g., the prediction of the observed dependence upon the luminosity and the relationship to the observed terminal velocities. In more recent work than is reported in this monograph, the theory has been used by Kudritzki and associates and by Abbott to predict the dependence of mass-loss rate and terminal velocity on the overall chemical composition. Recently reported observations of Magellanic Cloud O stars by Garmany and Conti confirm the predictions of lower terminal velocities in the SMC but smaller mass-loss rates are not found. A possible explanation might lie in an inadequate determination of the luminosities of the Magellanic Cloud stars due to uncertainties in the spectral type/effective temperature dependence. This important issue is currently unsettled.

A notable exception to the success of current radiatively driven stellar wind theory is observed stellar variability. This can only be understood in terms of the radiatively driven theory *if* the stellar radiation itself varies or *if* instabilities are naturally occurring in the wind. It would be important to demonstrate that the kinds of subtle details in the variability found in the stellar photospheres, as considered by Baade in Chapter 3, show up in a direct manner in the stellar winds. So far, there has been no quantitative work on this issue for O or Wolf-Rayet stars, but this would be a promising route to explore. The bottleneck is primarily an observational one that needs a dedicated and concentrated attack on several key stars at various wavelengths.

Stellar wind variability, particularly in the "narrow" components frequently found in the absorption part of the P Cygni profiles, is considered by Dr. Huib Henrichs in Chapter 4. After his discussion of the data, specifically the time scales, he considers various models that have been put forward to explain the existence of the narrow features. Most of these models cannot explain the data in any satisfactory way. He favors a connection to the NRP phenomena, in particular the mode switching, as a way of adding nonthermal energy to the radiatively driven wind. Although much more work, both observational and theoretical, will have to be addressed to this issue, many guess that it will turn out to be the right approach. Thus the winds of O (and Wolf-Rayet) stars are primarily radiatively driven, but additional energy which modifies the overall structure is coming from surface phenomena such as nonradial pulsations.

A star is a gaseous entity held in a delicate balance between the inward force of gravity and the outward force of radiation pressure due to the nuclear fires in its core. The "surface" or photosphere is the region from which the radiation escapes. But the star does not "end" at this point since matter still exists beyond it. In the stars of this monograph, *substantial* material is found above the photosphere in the form of a stellar wind. This wind has interactions with material further from the star either from its previous integrated history of mass loss, or from possible episodic "events," or from that matter initially found surrounding the star. This material, which is partially composed of stellar ejecta and partly of swept up interstellar material, has been substantially modified by the star itself, both by its intense radiation but also by its wind momentum. I have referred to this matter as the *Local* Stellar Medium, or LSM, to distinguish it from the relatively unaltered *Inter*stellar Medium, or ISM.

O and Wolf-Rayet stars greatly modify their environs; in star formation processes, the appearance of one or more newly born O stars dramatically affects further star births. Their deaths as supernovae will drastically affect the surroundings and inexorably alter what follows in the vicinity. In addition to these singular events, the study of the interaction of the star with its environment during its few million year lifetime is one of the most interesting of the newly emerging phenomena concerned with these hot stars. Many of the physical processes are nonlinear and are likely to be nonthermal as well. Most of the work done so far has been based on assuming spherical symmetry and a homogeneous medium. It is clear the ISM is highly nonhomogeneous; the LSM should be even more so since it has been modified by the stellar wind and the radiation field. Examples of the LSM are the so-called "rings" and "bubbles" associated with a few O and Wolf-Rayet stars. The conditions near other stars may not be favorable for easy detections of these features.

One of the main thrusts of the work described in Part II of this monograph is the presumption that Wolf-Rayet stars are the highly evolved descendants of the most massive stars. Both observations and theoretical modeling lead in the direction of this basic conclusion. Independent evidence comes from several avenues.

It is well known that massive stars are losing mass, given the ubiquitous presence of P Cygni profiles in the UV resonance lines; furthermore, surrounding several Of stars and some LBV is found nitrogen-enriched material, which must have been ejected from the star. It is reasonable to conclude that mass loss and stellar mixing must play a role in the evolution of massive stars, as they still contain hydrogen and are still burning hydrogen in their stellar interiors.

Red supergiant stars are known to be helium core-burning objects, yet they are not found brighter than a bolometric magnitude of about –10, corresponding to stars with initial masses around 40 $M_\odot$. Yet bolometrically brighter blue supergiants are known and stars more massive have been found; these supergiants must carry out their helium core-burning phases somewhere other than as red supergiants.

Wolf-Rayet stars are observed to have properties of anomalous composition of helium and nitrogen or helium, carbon, and oxygen, and very little or, in most cases, no hydrogen. Those stars that are members of binary systems are observed to be more luminous, yet less massive, than their companions. An assertion that the Wolf-Rayet stars are the helium burning descendants of the most massive stars fits all the previously observed data.

Independent confirmation of this picture can be seen in the overall galactic concentration of the most massive O stars, B supergiants, and the Wolf-Rayet stars in the spiral arms towards the galactic center; less massive O stars do not share such a longitude distribution and B main sequence stars and giants are even less similar in their locations.

Stellar models, with mass loss and mixing, are able to produce completely stripped cores in the helium-burning phases. These objects are found in the relatively hot part of the HR diagram, rather than in the cooler red supergiant portions. The enigmatic appearance of exclusively emission-line spectra can be reproduced from strong stellar winds, themselves the products of the extremes of luminosity and temperature one would expect of massive helium-burning objects with little or no hydrogen remaining. The interpretation of Wolf-Rayet stars as helium-burning objects is in accord with all the observed data and consistent with the expectations from stellar evolution theory.

The Of stars appear to be the late phases of the core hydrogen-burning phases of massive stars; stellar models suggest these stars may already be turning "leftwards" in the upper part of the HR diagram. Mass may also be lost in episodic events connected with the LBV stars, which potentially play an intermediate role between the H- and He-burning phases in some (all?) massive stars. This connection is not yet certain. Binaries probably play only a relatively minor role in massive star evolution unless the companions are initially very close since the models indicate the stars turn back to the "left" while still burning hydrogen before their radii become appreciable.

The types and numbers of the various Wolf-Rayet stars are found to be different from place to place, but these relationships are not yet sorted out due to the difficulty in predicting the effective temperatures from the stellar models. Potentially, these stars have the ability to give us information concerning the evolutionary details of the most massive stars in different environments.

Epilogue to Part II

The contributions of Part II of this monograph were essentially completed by the end of 1986. Early in 1987, the radiation from the supernova explosion in the LMC reached the Earth, giving us an unprecedented view of the violent death of a massive star. The immediate progenitor object, NS-69 202, was a B3 supergiant; there is circumstantial evidence of a prior history as a red supergiant. The initial mass of the object inferred from the bolometric luminosity of the explosion was about 20 $M_\odot$, roughly O9 on the main sequence.

Our prior knowledge of stellar evolution would have suggested that such a 20 $M_\odot$ star would end its life in the red supergiant stage. For some reasons not yet fully understood, NS-69 202 returned to the blue part of the HR diagram before the explosion. As a blue supergiant, it had a relatively small radius and thus the optical light was rather faint compared to other SNII. After the event, it has been possible to parameterize mass loss and mixing in a series of 20 $M_\odot$ evolving models in which this scenario might normally occur under similar initial conditions. One would then predict that other relatively faint SN of type II, like 1987a, would be common. So far, SNII have not been identified in other galaxies but possibly a selection effect has been operating against detecting them.

The shock wave from the SN event is now proceeding outward through the LSM of the exploded star. As the shock wave proceeds, it encounters and excites material that had been previously ejected by the prior blue and red supergiant phases which may be distinguished from one another by their different "terminal" velocities. Observations over time will thus reveal some of the past history of NS-69 202, in particular, its composition. It would not be too surprising to find enhanced nitrogen in these stages; the detection of anomalous helium will probably be more difficult, given the physical conditions of the gas cloud as the shock reaches it. SN 1987a will certainly be a revealing object over the next years as the physical conditions earlier and earlier in its history are brought to light. As the material expands from the SN, it becomes less opaque, revealing the innermost core of the star in the times immediately preceding the explosion. New and probably unprecedented information awaits us in the gradual unfolding of a massive star. Our understanding of such massive stars may well be completely altered by such observations and interpretation.

Part III

In Part III, it is shown that the strengths of the emission lines and of almost all of the strong absorption lines in the spectra of O and Wolf-Rayet stars, the emission of X rays and an infrared and radio excess, and the outflow of wind cannot be interpreted in a self-consistent way if one uses only classical stellar atmosphere theory based on the constraints of hydrostatic, radiative, and statistical equilibrium. The general problem of understanding the spectra of O, Of, and Wolf-Rayet stars is posed in Chapter 6. In Chapter 7, the theory of the atmospheres of hot stars published by the end of 1981 is reviewed and its successes and its failures are noted. In Chapter 8, a study is made of the effect of assuming a relatively weak, inhomogeneous magnetic field in addition to the assumptions upon which the classical theory of stellar spectra has long been based. It is argued that this additional assumption contributes greatly to understanding what is observed.

Radiatively driven turbulence may be the cause of some of the spectroscopic phenomena which are observed, but it seems most likely that the prime causes of the high temperatures and outflow deduced to be present in the mantles of O, Of, and Wolf-Rayet stars are the heating and momentum transfer which result from turbulent motions in the photosphere buffeting the footpoints of magnetic lines of force which thread the photosphere. In Chapter 8, it is postulated that the needed magnetic lines of force originate in dynamos situated in the He^+ convective layer which occurs in the deepest parts of the photospheres of O and Wolf-Rayet stars. The primary source of the motion described as turbulence and invoked to generate magnetohydrodynamic (MHD) waves (which then propagate through the mantles of the O, Of, and Wolf-Rayet stars) is the rotation of the star. The viscosity of the rotating fluid which forms the boundary layer of the star is believed to generate a turbulent, helical flow in the photosphere, a flow which is parallel to the surface of the star. This flow then creates dynamos in the He^+ convection zone and generates MHD waves.

Radiatively driven turbulence (see Chapter 7) may be generated when radiation traverses a low-density ionized gas in which there is a velocity gradient. However, it remains to be shown that the sound waves needed in the theories of radiatively driven turbulence can propagate from the photosphere, or deeper layers where they may be generated, to the mantle of the star. The necessary velocity gradient for the theories of radiatively driven turbulence can be generated by the action of MHD waves (see Chapter 8). Once a significant velocity gradient has been generated, the effects of radiation pressure in lines may augment the velocity gradient generated by the action of MHD waves.

Five questions are posed at the end of Chapter 6. Brief answers to these questions are as follows:

1. The ultimate source for the nonradiative energy and momentum which are released and which cause a mantle to form is the rotation of the star.

2. The energy and momentum which originate in the rotation of the star are transferred to the gas in the mantle of the star by the actions of MHD waves as they propagate along the magnetic field lines which traverse the mantle. The magnetic field lines are generated in dynamos which arise as helical motion, generated by the rotation of the star in the weak general magnetic field. It is likely that the magnetic field lines pass through the photosphere only in narrow regions concentrated at the edges of turbulent eddies, as is observed in the Sun.

3. The release of nonradiative energy and momentum by the actions of MHD waves is not significant in the photosphere because, in the parts of the photosphere where magnetic field lines are present, the plasma beta (ratio of gas pressure to magnetic pressure) is large. In the parts of the photosphere free from magnetic field lines, no MHD waves are present.

4. Nonradiative phenomena in stellar atmospheres are important for understanding the evolution of stars in all cases where the empirically selected criteria for spectral type are spectroscopic features dependent for their strength and appearance on the presence of nonradiatively generated heat and momentum in the mantle. It is commonly assumed that a spectral type may always be used as an unambiguous indicator of the primary parameters of the theory of stellar evolution. We recall that the primary parameters of the theory of stellar evolution are the mass, radius, effective temperature, and composition of the star. It has been shown in Chapter 7 that these basic parameters cannot be deduced unambiguously from the spectra of O, Of, and Wolf-Rayet stars when one solely uses classical theory. The observed appearance of O, Of, and Wolf-Rayet spectra depends sensitively on the amount of nonradiative energy and momentum in the mantle of the star as well as on the size of the mantle and the state of motion. Consequently, it is very important to take note of the occurrence of nonradiative phenomena in the atmospheres of hot stars if one is to understand the evolution of the hot stars correctly. Information about the masses, radii, and effective temperatures of O and Wolf-Rayet stars may be found in Part I of this book. The theoretical studies reviewed in Part III indicate that Population I O, Of, and Wolf-Rayet stars have normal, solar composition. In some subluminous O stars, the relative abundance of helium to hydrogen is not normal. It appears that gravitational settling of helium may have taken place in some cases, while in others most of the hydrogen-rich envelope may have been lost.

5. The observation of nonradiative phenomena in the atmospheres of O, Of, and Wolf-Rayet stars do not give guidance concerning the stage of evolution of isolated, single, O, Of, and Wolf-Rayet stars. Spectra which are classified as O, Of, and Wolf-Rayet are found among the normal, massive, young stars of Population I and among the subluminous stars. The latter stars can rightfully be regarded as old, highly evolved objects. The information presented in Part III of this book indicates that the spectroscopic phenomena which typify O, Of, and Wolf-Rayet spectra are most likely due to the heating and momentum transfer in a mantle caused by the rotation of a hot star which has a weak magnetic field in its boundary layer. The boundary-layer magnetic field may be primordial in origin, or it may have been generated during the evolution of the star. Until the theory of stellar evolution reaches such a state of completeness that one can predict the occurrence of magnetic flux in the boundary layer of a hot star (e.g., specifically in the He^+ convection zone of its photosphere), one cannot only use the observation of spectroscopic phenomena which suggest heating and momentum transfer primarily generated by the actions of MHD waves in the mantle of the star to deduce the stage of evolution of a hot star. Other independent information such as the composition of the atmosphere, the luminosity of the star, its effective temperature, and its mass are needed before one may estimate the stage of evolution of the star.

In summary, the appearance of the spectra of O, Of, and Wolf- Rayet stars is strongly conditioned by the actions of nonradiative heating and nonradiative momentum transfer in the mantle of the star. Therefore, an O, Of, or Wolf-Rayet spectral type is not an unambiguous guide to the stage of evolution of the star.

It may be possible to deduce interesting evolutionary information from spectral types of classes of O, Of, and Wolf-Rayet stars when the theory of stellar evolution has been extended to include a prediction of when in their lifetimes massive model stars (and those of low mass), having high effective temperatures, may have magnetic flux in their subsurface He^+ convection zones and when they may be rotating at such a speed that helical flow motion is generated in the boundary layer of the star. The latter flow will create subsurface dynamos in the general magnetic field with a consequent expulsion of magnetic lines of force from the star and the formation of a mantle having the properties which are deduced for the mantles of O, Of, and Wolf-Rayet stars. In Chapter 8, it is explained how the expulsion of magnetic lines of force can account for the mantles of O, Of, and Wolf-Rayet stars including the rate of mass loss from each star.

1

INTRODUCING THE O AND WOLF-RAYET STARS

I. INTRODUCTION

The aim of Part I of this book is to give some general information on O and Wolf-Rayet stars—how they are defined and what their main observable properties are.

The O and Wolf-Rayet stars are defined by the appearance of their spectra in the visual spectral range which is observable from the ground. The O-type spectra have a hot energy distribution and absorption lines of He II, He I, and H I. They can be ordered in sequences by using a choice of changing observable details. Theoretical model atmospheres show that, with an adequate choice of the changing details, the sequences can represent either increasing effective temperature or increasing surface gravity. Some spectroscopic features originating in outer parts of the stellar atmosphere are not well correlated with effective temperature or surface gravity. These features are generally avoided in normal classification schemes as described in Part I, and the problem of modeling the outer parts of the stellar atmospheres is treated in Part III. Although O stars are defined by spectroscopic features which correlate well with effective temperature or surface gravity, they *do not form an homogeneous group*, and stars with very similar spectra may differ widely in their masses, radii, ages, internal structure, etc. Most of the O stars situated near the galactic plane are bright young massive objects. Their internal structure and evolution seem to be well understood. They form an extension to higher temperatures of the B-type stars and are generally called "normal O stars." In the halo, a large number of intrinsically faint blue stars have been discovered, with energy distributions similar to the energy distribution of normal O or B stars. They are called "O subdwarfs" when He II absorption lines are present and "B subdwarfs" when He II is absent. With their hot atmospheres and low luminosities, the subdwarfs have small masses and can correspond only to late stages of evolution. These late stages of evolution are not yet completely understood, and the relation between O and B subdwarfs is still unclear. A puzzling point is that the atmosphere of an O subdwarf can apparently mimic almost exactly a normal O star. The same spectroscopic features may thus correspond to very different objects.

The Wolf-Rayet spectra are dominated by strong and broad emission lines. They can be and are also ordered in sequences, but there is no evidence that these sequences correspond to a monotonic change in effective temperature or in any other physical parameter.

The information summarized here is, on the whole, that available by the early eighties, supplemented by a few later results. Some of the topics treated in this Part of the book are expanded in Parts II and III in order to take into account recent observations and theories.* In particular, a rapid increase of knowledge about Wolf-Rayet stars is occurring at this time.

II. SPECTRAL CLASSIFICATION

The classification systems used for O stars are based either on the line spectrum or on the continuum. Here we review the criteria used to select O and Wolf-Rayet stars.

A. The First Spectral Classification Systems

The early history of spectral classification has been summarized by R. H. Curtis (1932), and the classification systems for O stars which were developed before 1965 have been discussed by Underhill (1966). In the most-used system—the MK system which was originated by Morgan and Keenan in 1943—the spectral types are based solely on the apparent relative intensities of lines of He II and He I in the spectral region between 4000 and 5000 Å. Typical well-widened spectra obtained with grating spectrographs at ~125 Å mm^{-1} (see Morgan et al., 1978) show that the ratios He I 4471/He II 4541 and He I, He II 4026/He II 4200 decrease monotonically from type B0V to O4 owing to the decrease in strength of He I 4471 and of the He I/He II blend at 4026 Å and to the increase in strength of He II 4541 and He II 4200 as one progresses from type B0V through O9, O8, etc. to O4. The He II lines are members of the Pickering series due to transitions from the levels with chief quantum number equal 4 to higher levels. The He I lines are members of the 2^3P^o–n^3D series.

The standard stars which define the MK classification system for the O stars have been given by Johnson and Morgan (1953) and by Morgan et al. (1978). Interpolation between the spectra of these stars, by means of visual inspection of well-widened spectra covering the range from 3900 to 5000 Å permits classification into the various O subtypes. The MK classification system for the O stars does not provide a separation into luminosity classes. The spectral types run from O9.5, which follows type B0, by half steps, to O4. In other parts of the HR diagram, the MK luminosity classes are used to provide a two-dimensional separation into main-sequence stars, which form classes V and IV, into giants, which form classes III and II, and into supergiants, which form classes Ib, Iab, and Ia.

Many O-type stars show emission lines at N III 4634, 40, 41 and at He II 4686. Pearce (1930) introduced the designation "f" to indicate those O stars which showed emission at the N III multiplet and at He II 4686 on prismatic spectra having a dispersion of about 51 Å mm^{-1} at Hγ. Some of these stars also show emission at Hα, C III 5696 and at other lines. Pearce chose not to use the letter "e" to indicate the presence of these emission lines because the type Oe5 had been used in the *Henry Draper Catalogue* to identify O stars which showed an *absorption-line* spectrum on the objective-prism plates used for that first comprehensive systematic survey of the spectral types of stars. The Victoria system of classification for the O stars is based primarily on the relative strengths of the He II and the He I lines because simple theory indicates that this should be a criterion responsive to effective temperature.

B. Revisions Derived from Low-Dispersion Spectra

A number of weak absorption lines due chiefly to ions of C, N, O, and Si can be seen at types B2 through O—the OB group—on spectra having a dispersion of 100 Å mm^{-1} or better. Walborn (1971a, 1972, 1973a) has used the appearance of the spectra of the OB stars at a dispersion of *62 or 63 Å mm^{-1}* covering the region 3900 to 4800 Å to develop a more detailed spectral classification of the hottest absorption-line stars than is possible using the low-dispersion spectra of the MK system. His system assigns luminosity classes to these stars. He follows the method of the MK system and

*Especially for effective temperatures; see Chapter 3.

defines his subtypes by means of standard stars. The various subtypes in the range B2.5 to O9 are based chiefly on criteria which involve lines from He I and He II and from Si II, Si III, and Si IV. Those in the range O8.5 to O3 are based on the ratio He II 4541/He I 4471 for spectral type and on a variety of criteria for luminosity class, including the presence of emission lines of N III and He II. Walborn's primary classification criteria are listed in Table 1-1. His standard stars, published in Walborn (1971a, 1972, 1973a), are given in Table 1-2. In these three papers, Walborn gives the spectral type for most of the bright O stars and for some stars with types in the range B2.5 to B0.

Walborn retains use of the letter "n" (meaning nebulous or diffuse lines) as an indication of the width of the absorption lines seen on his spectra. He uses ((n)) to indicate that Si IV 4116 and He I 4121 are just merged at 63 Å mm^{-1}, (n) to represent a degree of broadening intermediate between that of ((n)) and the broadening seen in the stars which are given n-types at 125 Å mm^{-1} (the typical dispersion for the MK system), and [n] to indicate that the broadening of the H lines is much greater than that of the He lines. These designations are not surely related to rotational broadening; in some cases, they refer to double-lined spectra which were not resolved at 63 Å mm^{-1}.

Walborn also retains use of the letter "f" and expands this notation to designate various cases which are seen. He uses ((f)) to indicate weak N III 4634,40,41 emission coupled with strong He II 4686 absorption, (f) to indicate the presence of N III 4634,40,41 emission and He II 4686 absorption which is weak or neutralized by being filled in with emission, f to indicate that both N III 4634,40,41 and He II 4686 are in emission, and f^+ to indicate that Si IV 4089 and 4116 lines are in emission, as well as the N III 4634,40,41 multiplet and He II 4686. The f^+ characteristic generally appears only in O6 to O4 stars.

The emission lines used to assign the f characteristics typically do not have a P-Cygni profile composed of a shortward displaced absorption component and a more or less undisplaced emission component, but appear to rise from the continuum with little or no underlying absorption visible.

The type O3f* (Walborn, 1971b, 1973b, 1982) is described as follows: He I is not visible; the N IV emission at 4058 is stronger than the N III emission at 4634,40,41; the N V absorption at 4604 and 4620 is comparable in strength to He II 4541; the lines Si IV 4089 and 4116 are weakly in emission; and the He II 4686 emission is shifted longward from its normal position. This type of spectrum is rather similar to that shown by the WN7-A and WN6-A types described by Hiltner and Schild (1966). (See below.) Walborn (1982) introduces also the designation (f*) and ((f*)) corresponding to weak or strong He II 4686 absorption.

Walborn (1973a) has noted two groups of O-type spectra which probably give evidence of different kinds of shell activity, giving them the notations "Of?p" and "Onfp." The prototypes of the first group are HD 108 and HD 148937, which are described in Walborn (1972). A primary characteristic is that emission in C III 4647,50,51 is comparable in strength with the emission in N III 4634,40,41. Profiles of the hydrogen lines usually have a marked P-Cygni character in this type of star. Study at higher dispersion indicates that this type of spectrum may be due to a binary system enveloped in gas. *Weak* C III 4647,50,51 emission is known to occur in a number of Of stars, but these stars are not considered to belong to the Of?p group because they do not show a complex emission- and absorption-line structure for the hydrogen lines.

The principal distinguishing characteristic of the Onfp stars is a composite emission and absorption structure at He II 4686. The appearance of the profile at 63 Å mm^{-1} may be described either as an absorption line with emission wings or as a relatively broad emission feature superposed on an absorption line. There is a range in profiles from strong absorption with very weak emission wings (HD 14434, HD 192281) to strong emission split by a weak absorption feature as in λ Cephei. This broad emission feature appears to be similar to those

Table 1-1
Primary Classification Criteria for OB Stars According to Walborn

Criterion	Used For	Notes
Types B2.5 — B1.5		
Appearance He I 4009	Luminosity class	Diffuse at class V
He I 4009/He I 4026	Spectral type	Maximum at B2
He I 4121/He I 4144	Luminosity class	
Si II 4128,30/He I 4121	Spectral type	Defines B2.5
Si III 4552/He I 4387	Luminosity class	
C III 4647,50,51/He I 4713	Spectral type	Assists in resolving diagonal ambiguity in B1 V — B2 III and in defining B1.5; use with caution
Types B1 — O9		
He I 4009/He I 4026	Spectral type	
Si IV 4089/Hδ or He I 4121	Spectral type and luminosity class	
Si IV 4116/He I 4121	Spectral type and luminosity class	Generally blended at MK dispersion
He II 4200/He I 4144	Spectral type	Defines O9, O9.5
He II 4541/He I 4471	Spectral type	Defines O9, O9.5
He II 4541/He I 4387	Luminosity class	Important at O8
Si III 4552/He II 4541	Spectral type	Defines O9.7
Si III 4552/He I 4387	Luminosity class	
Si III 4552/Si IV 4089	Spectral type	
He II 4686/He I 4713	Luminosity class	Negative luminosity effect
Types O8.5 — O3		
He II 4200/He I, II 4026	Spectral type	Equals unity at O6
He I 4471	Spectral type	Almost invisible at O4; invisible at O3
He II 4541/He I 4471	Spectral type	Equals unity at O7
Types O8.5 — O7		
Of; Si IV very strong	Luminosity class Ia	HD 151804, O8 Iaf
He II 4686 absorption weak or neutralized; Si IV strong	Luminosity class Ib	HD 225160, O8 Ib(f)
He II 4686 absorption; Si IV strong; He II 4541/He I 4387 greater than at V	Luminosity class III	ξ Per, O7.5 III((f)) λ Ori, O8 III((f))
He II 4686 absorption very strong	Luminosity class V	15 Mon, O7 V((f)) HD 46149, O8.5 V
Types O6 — O4		
Of and Of^{+}	Luminosity class I	HD 15570, O4 If^{+}
He II 4686 absorption weak or neutralized	Luminosity class III	HD 15558, O5 III(f)
He II 4686 absorption strong	Luminosity class V	HD 46223, O4 V((f)) HD 46150, O5 V((f))

Table 1-2
Classification Standard Stars for OB Types According to Walborn

Spectral Type	Dwarfs	Giants	Supergiants
B3	η Aur V, η UMa V	HD 21483 III	o^2 CMa Ia, κ Cru Ia
B2.5	HD 214432 V	π^2 Cyg III	3 Gem Ib, 55 Cyg Ia, HD 92964 Ia
B2	HD 42401 V	γ Ori III, HD 141318 II	9 Cep Ib, χ^2 Ori Ia
B1.5	HD 154445 V	12 Lac III, HD 96159 II	HD 190603 Ia$^+$
B1	ω^1 Sco V	σ Sco III, -57°3506A II	ζ Per Ib, HD 86606 Ib, HD 13854 Iab, HD 148688 Ia
B0.7	HD 201795 V	ε Per III	HD 190919 Ib, HD 109867 Ib, κ Cas Ia†, HD 152235 Ia
B0.5	HD 36960 V	1 Cas III	κ Ori Ia, HD 152234 Ia
B0.2	τ Sco V	HD 6675 III, HD 108639 III	–
B0	υ Ori V	HD 48434 III, HD 150041 III	ε Ori Ia, HD 91969 Ia
O9.7	–	HD 10125 II	ζ Ori Ib, HD 152147 Ib, μ Nor Iab, HD 152003 Iab, HD 195592 Ia
O9.5	AE Aur V, HD 93027 V	HD 189957 III	19 Cep Ib, α Cam Ia
OC9.5	–	–	HD 152249 Iab†
O9	10 Lac V, HD 93028 V	ι Ori III	HD 210809 Iab, HD 149404 Ia
O8	–	λ Ori III((f))	HD 225160 Ib(f), HD 151804 Iaf
O7.5	–	ξ Per III(n)((f))	–
O7	15 Mon V((f)), HD 91824 V((f))	HD 93222 III((f))	–
O6.5	HD 165052 V(n)((f))	HD 190864 III(f)	HD 69464 Ib(f)
O6	HD 101190 V((f))	–	λ Cep I(n)fp
O5	HD 46150 V((f)), HD 93204 V((f))	HD 15558 III(f)	–
O4	HD 46223 V((f)), HD 96715 V((f))	–	HD 190429A If$^+$
O3	HDE 303308 V((f))	–	VI Cyg No. 7 If*, HD 93129A If*

†Nitrogen-deficient star.

noted by Wilson (1955, 1957, 1958) and confirmed by Underhill (1958). The broad emission may be interpreted as coming from an extensive optically thin atmosphere, expanding or rotating, which surrounds the O star. (See Part III.) Walborn notes that the uniform presence of broadened absorption lines within the Onfp category is suggestive of rapid rotation. Either the n or the f characteristic, or both, may be in parentheses in this category of spectral type.

Walborn (1971c) has drawn attention to the fact that, in some stars of the OB group, absorption lines of N III or N II are enhanced in strength compared to those in normal stars of equivalent spectral type, while lines of carbon and/or oxygen may be weakened. He denotes the types for these stars by ON or BN, depending on whether the star is of type O or type B. He also introduces the symbols "OC" and "BC" to denote spectra in which the lines of N III or N II are weak compared to those in normal spectra of an equivalent type and the absorption lines of C III or C II are enhanced in some B-type supergiants. Walborn (1976) has listed 20 known OBN and OBC stars. Here he discusses the basis for the classification and possible origins for the observed anomalies. He also lists stars which have been associated with the OBN/OBC group but are normal and 21 stars with moderate CNO anomalies. The calculations of Dufton and Hibbert (1981) show that certain N II lines, including the line at 3995 Å, can be greatly strengthened by statistical equilibrium (non-LTE) effects in the atmosphere.

The interpretation of the discrepancies noted by Walborn in the relative intensities of various lines in O- and B-type spectra and codified by specific spectral-type notations is not straightforward. Many of the lines that Walborn uses for spectral-type criteria are affected by non-LTE effects or by emission of a selective character. (See Part III.) Thus, caution is desirable in interpreting Walborn's spectral types and luminosity classes simply in terms of effective temperature and gravity in the stellar atmosphere and of concluding that anomalous line strength means anomalous abundance. The calibration of Walborn's luminosity classes in terms of visual absolute magnitude is discussed in the section *Absolute Magnitudes*.

C. Revision Using Moderate-Dispersion Spectra

Conti and his colleagues have studied O-type spectra on spectrograms of moderate dispersion, and Conti has developed a classification scheme based on the equivalent widths of He I 4471 and He II 4541 measured on spectra having a dispersion of *16 Å mm*$^{-1}$. His criteria, presented in Conti and Alschuler (1971) and Conti and Frost (1977), are given in Table 1-3. The measured variable is:

$$\log_{10} W' = \log_{10} W(4471) - \log_{10} W(4541), \quad (1\text{-}1)$$

where W is the equivalent width of the indicated line determined from spectra having a dispersion of 16 Å mm^{-1}. The most recent spectral types assigned by Conti for 150 O stars are given by Conti and Leep (1974) and by Conti and Frost (1977). Eighteen of the stars listed by Conti and Frost are in the list of Conti and Leep or that of Conti and Burnichon (1975), sometimes with differing types.

Conti's spectral types correlate well with those of Walborn. He has adopted Walborn's convention with regard to stars with abnormally strong N III or C III lines (i.e., he designates such spectra as ON or OC, respectively). He also adopts Walborn's convention of using ((f)) to denote stars with N III 4634,40,41 in emission but He II 4686 in absorption, (f) for those with N III 4634,40,41 emission but He II 4686 either very weak in absorption or completely neutralized, and 'f' for those with both N III 4634,40,41 and He II 4686 in emission above the level of the continuum. He does not adopt Walborn's convention about use of the letter "n" to indicate the apparent width of the absorption lines.

Table 1-3
Criteria Used by Conti to Classify O Stars

Spectral Type	Range for log W'	Spectral Type	Range for log W'
O9.5	$\geqslant +0.45$	O6.5	-0.11 to -0.20
O9	$+0.44$ to $+0.30$	O6	-0.21 to -0.30
O8.5	$+0.29$ to $+0.20$	O5.5	-0.31 to -0.45
O8	$+0.19$ to $+0.10$	O5	-0.46 to -0.60
O7.5	$+0.09$ to $+0.00$	O4	<-0.60
O7	-0.01 to -0.10	O3	He I 4471 absent

Luminosity classes V, III, or I are assigned by Conti for most stars of spectral types O7 and later on the basis of the value of the ratio Si IV 4089/He I 4143, but Conti has not published a quantitative relationship between the value of this ratio and luminosity class or visual absolute magnitude. An analysis of Conti's measurements of the strength of He II 4686 in O stars and the visual absolute magnitudes of these stars as determined from membership in open clusters and associations suggests that the strength of He II 4686 is related to luminosity (Conti and Leep, 1974). Likewise, all the most luminous O stars have the f characteristic. However, Conti does not attempt to assign luminosity classes for stars of types O6 and earlier. There is a spread of only about 2 mag in visual absolute magnitude for these stars. The calibration of Conti's types in terms of visual absolute magnitude is discussed in the section *Absolute Magnitudes*.

Conti reserves use of the letter "p" for stars whose spectrum cannot be interpreted in terms of "normal," or expanding envelopes. Thus, he does not follow Walborn in assigning the letter "p" to spectra which contain lines with a typical P-Cygni profile consisting of a shortward displaced absorption component and a more or less undisplaced emission component. He adopts a particular convention with regard to the use of the letter "e" which is not used by Walborn. Conti and Leep (1974) have defined Oe stars as those O stars with emission in the H lines, at least Hα, but *no* emission in N III or other lines. The type Oe isolates a small class of stars related to the Be stars. The designation, O(e), is used when the effects are mild. The Oe and O(e) stars have absorption lines which are broad and suggestive of large rotational velocities; frequently, there is a central absorption dip in the emission feature. These spectra are like Be spectra except that they have He II 4541 in absorption. All are in luminosity classes V or III. The designation Oef is used when the stars show an emission structure in He II 4686 that is like that in the H lines for Oe stars and N III 4634,40,41 is in emission. The Oef and O(ef) stars are all of early spectral type, the latest type being O6.5; the Oe stars are generally of late O type. Conti's type Oef is about equivalent to Walborn's type Onfp.

D. BCD Spectral Classification

This quantitative spectral classification scheme for O, B, A, and F stars was initiated by Barbier and Chalonge (1941) and developed by Chalonge and Divan (1952, 1973). It is based on the continuous spectrum, and the classification parameters used are: (1) the size of the Balmer jump D (where $D = \log F_R - \log F_V$, F_R and F_V being the fluxes on the longward side and on the shortward side of the Balmer jump, respectively), and (2) the effective wavelength λ_1 of the Balmer jump as seen on the low-dispersion prismatic spectra used by Chalonge and Divan (220 Å mm^{-1} at Hγ).

Most of the O stars are more or less heavily obscured by interstellar matter, but the measurement of λ_1 and the Balmer jump *as done* in the BCD system is not perturbed by the presence of interstellar extinction. Since these parameters characterize the physical conditions of hydrogen in the photosphere, they are relatively insensitive to chemical anomalies or to the presence of low-density circumstellar material and are therefore well correlated with model parameters such as effective temperature and gravity. The Balmer discontinuity of the model atmospheres corresponding to O stars, as well as that of real stars, depends mostly on the effective temperature. The wavelength λ_1 is very sensitive to the gravity through the influence of the Stark effect on the Balmer line wings: it is easy to see in Figure 1-1 that λ_1 increases when the Balmer line wings are broader. Stellar rotation and chromospheric phenomena affect mainly the cores of the Balmer lines; they have almost no influence on the value of λ_1,

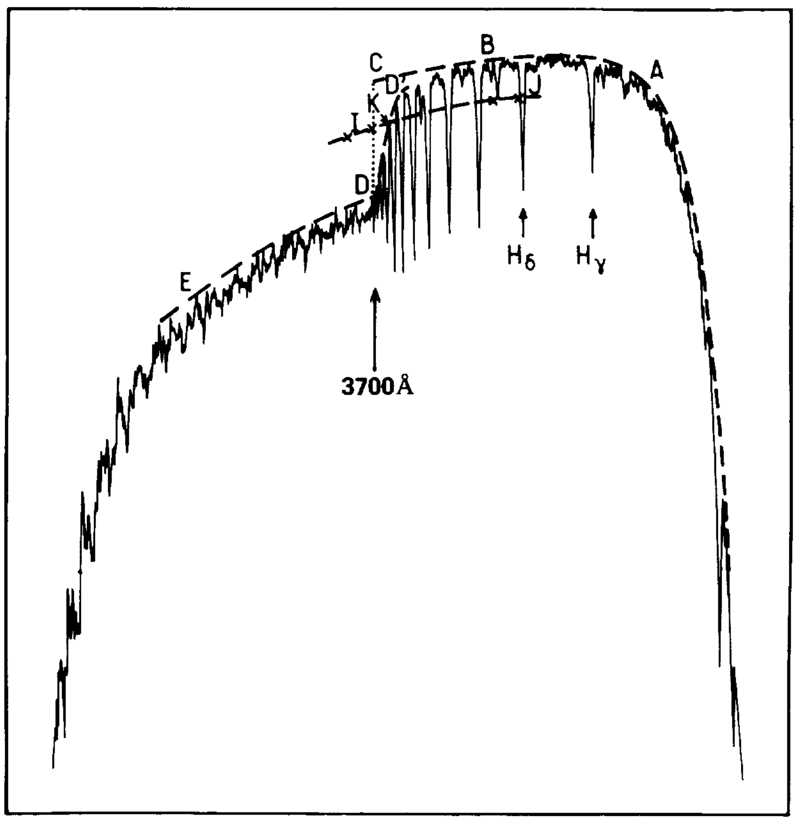


Figure 1-1. The position of the Balmer discontinuity as defined by the parameter λ_l. Dashed lines AB and DE are the Paschen and Balmer continua, respectively, and DD' is a smooth curve joining the points of highest density between the Balmer lines. Curve IJ is deduced from the Paschen and Balmer continua by subtracting D/2 from the intensities which correspond to curve AB and by adding D/2 to those corresponding to curve DE, D being the value of the Balmer discontinuity obtained by means of the spectrophotometric measurements. Parameter λ_l is the wavelength corresponding to the point, K, where the curves DD' and IJ intersect. In the case of O stars, λ_l increases from about 3740 Å for supergiants to about 3770 Å for main-sequence stars (see Figure 1-2).

which is thus a purely photospheric parameter. The BCD classification system and the classifications based on the line spectrum are closely correlated, and the parameters D and λ_1 have been calibrated in terms of MK spectral types and luminosity classes. For O stars, D is a useful parameter for spectral type although it depends also on luminosity class (see Figure 1-2), while the value of λ_1 is a good criterion for luminosity class. The parameters, D and λ_1, have also been calibrated in terms of absolute magnitudes (see Figure 1-2). More details are given in *B Stars With and Without Emission Lines* (Underhill and Doazan, 1982).

The facts that for O stars the Balmer jump is very small and that the uncertainty in λ_1 is much larger than in the case of the B or A stars render the classification of O stars by the BCD method somewhat difficult. However, the use of new fine-grain photographic plates greatly improves the precision in λ_1 from what was originally possible and permits a secure classification of O stars in the BCD system. In Figure 1-2, the limits of the MK spectral-type/luminosity class boxes and the curves of equal absolute magnitude in the $\lambda_1 D$ plane are from Chalonge and Divan (1973). The positions of the circles and squares correspond to new values of λ_1 and D (Divan, unpublished) obtained for sixteen O stars that had been classified by Walborn (1973a) and/or by Conti and Leep (1974). The luminosity classes inscribed inside the circles and squares are from Walborn in the upper diagram and from Conti and Leep in the lower diagram. Twelve stars (those represented by the circles) are common to the two diagrams. The stars represented by squares were classified only by Walborn or only

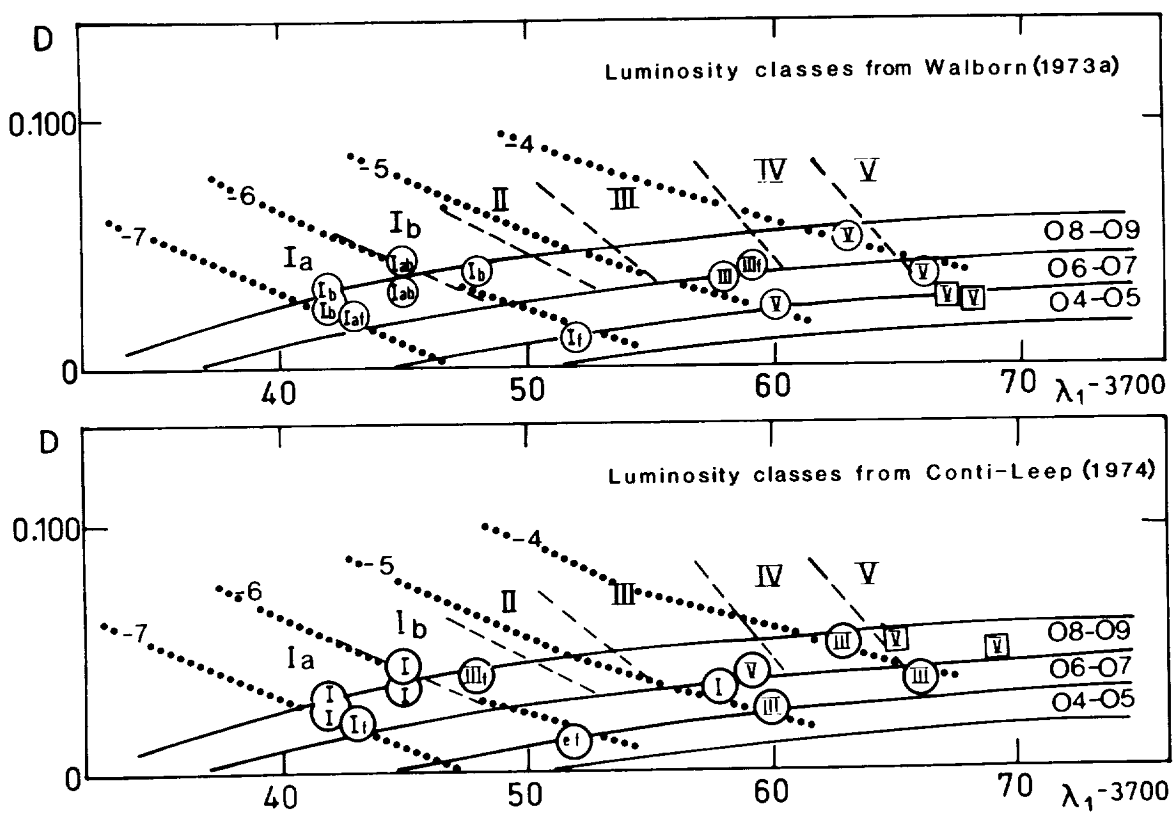


Figure 1-2. The BCD classification scheme for O stars of Chalonge and Divan. Here the limits of the MK spectral-type/luminosity class boxes are defined by the dashed and solid lines. The curves of equal absolute magnitude in the λ_1 D plane for O stars (Chalonge and Divan, 1973) are given by dotted lines. For the meaning of circles and squares, see the text.

by Conti and Leep. These diagrams show that the $\lambda_1 D$ luminosity classes are in good agreement with classes assigned from the line spectra of O stars. The agreement is particularly good with the luminosity classes of Walborn.

All the classification schemes described in these first paragraphs apply to the so-called normal O stars. We shall now give some indications on the classification of O subdwarfs and Wolf-Rayet stars.

E. Classification of O Subdwarfs

The O and B subdwarfs, although discovered as a group of statistically low-luminosity stars in proper motion surveys, are now defined by spectroscopic features which give the possibility of recognizing the O or B subdwarf character in individual stars. The O and B subdwarfs have a hot energy distribution and broader lines than those in the corresponding main-sequence stars, the Balmer series being visible only up to $n = 10$ or 12. The star is called an O subdwarf if the line He II λ4686 is visible, and a B subdwarf if He II λ4686 is not visible.

However, no classification scheme is available for O subdwarfs. The normal O stars can be ordered in a two-parameter classification scheme because of their uniform chemical composition, the observed spectrum depending only on the effective temperature and the surface gravity. Detailed analyses of O subdwarf spectra with the aid of adequate model atmospheres such as those described in the section *Effective Temperatures* show that the atmospheric chemical composition of O subdwarfs spans a large range from almost pure hydrogen to almost pure helium. Any classification system for these stars should then be three-dimensional. The parameters related to effective temperature or to surface gravity used in normal O-star classifications are sensitive to the chemical composition and have no meaning in the case of O subdwarfs. At the present time, the atmospheric parameters of an O subdwarf can be derived only from a complete analysis of the spectrum. Even the significance of the adopted limit between O and B subdwarfs is unclear, the presence or absence of He II 4686 depending on the chemical composition and on the gravity, as well as on the effective temperature (see, for instance, Kudritzki and Simon, 1978).

F. Classification of Wolf-Rayet Stars

The words "Wolf-Rayet" refer to a particular type of spectrum which Wolf and Rayet (1867) were the first to recognize as they surveyed the spectra of stars in Cygnus with a visual spectroscope. In place of the usual continuous spectrum crossed by rather narrow absorption lines, Wolf and Rayet saw a few relatively broad emission lines superposed on a faint continuous spectrum. Since 1867, about 160 Wolf-Rayet stars have been recognized in our Galaxy (see van der Hucht et al., 1981), 100 in the Large Magellanic Cloud (LMC) (Breysacher, 1981), 8 in the Small Magellanic Cloud (SMC) (Azzopardi and Breysacher, 1979), and about 80 in M33 (Wray and Corso, 1972; Corso, 1975; Massey and Conti, 1983b).

The spectral types of normal stars are based on absorption lines seen against a continuous spectrum. They can be interpreted grossly in terms of plane parallel layers of gas characterized by an effective temperature and a value for log g. (See Part III.) The effective temperature, with the constraint of radiative equilibrium, serves to determine the electron temperature in the atmospheric layers, while the acceleration of gravity, with the constraint of hydrostatic equilibrium, serves to determine the density in the atmosphere. A different absorption-line spectrum, thus the spectral type, is found for each pair of values of effective temperature and acceleration of gravity. Normal spectral types are so defined that they change monotonically with increasing effective temperature and with the luminosity of the star, which quantity relates monotonically to log g.

However, the classical model for a stellar atmosphere cannot explain a spectrum which is dominated by emission lines. Thus, although the set of spectral types assigned to Wolf-Rayet stars can be, and is, arranged in a reproducible order of changing detail, there is no sure evidence that the adopted types correspond to monotonic changes in effective temperature and log *g*. The spectral types corresponding to lines of the highest level of ionization, as seen in the spectral region between 3100 and 7000 Å, are given the smallest numbers in the hope that continuing the convention used for normal absorption-line stellar spectra will arrange the Wolf-Rayet stars in a sequence of monotonically increasing effective temperature. There are no spectroscopic criteria that can be uniquely related to log *g* for the Wolf-Rayet stars.

Two ways in which the classical model atmosphere may be changed to produce a model which can account for the presence of emission lines are: (1) to postulate an electron temperature in the outer parts of the stellar atmosphere which is higher than what is allowed by use of the constraint of radiative equilibrium and an assigned effective temperature, and (2) to postulate that the star appears to be much larger when it is seen in the wavelengths of strong emission lines than when it is seen in wavelengths of the continuous spectrum. The development of these ideas is discussed further in Part III. The general intensity distribution observed in the continuous spectra of Wolf-Rayet stars is similar to that found for O stars.

A Wolf-Rayet spectrum, whatever its source, has been defined to have the following characteristics (Thomas, 1968):

1. The spectrum consists of emission lines on a continuous spectrum which has an energy distribution rather like that of an O or a B star.
2. A few absorption lines may occur as shortward displaced satellites on the edges of some of the emission lines, but there is no general absorption line spectrum as known for normal spectral types. If such a spectrum is seen, it is attributed to a companion star. However, recent observations (Moffat and Seggewiss, 1978; Conti et al., 1979) suggest that intrinsic absorption lines occur for some WN7 stars.
3. The emission lines seen in any one object represent a wide range of excitation and ionization, the excitation of the line spectrum generally being higher than that estimated from the shape of the continuous spectrum.
4. Most of the emission lines are broad with full width at half maximum (FWHM) from several hundred to several thousand km s^{-1}. The widths are not the same for all lines in any one stellar spectrum.
5. Most of the spectra fall into two groups: (1) the WC stars in which the lines from ions of C and O dominate, and (2) the WN stars in which the lines from ions of N dominate. Both groups show strong lines of He II.

The following paragraphs summarize the empirical descriptions which have been adopted for the Wolf-Rayet spectral types. The spectrum of each star differs in some respects from those of all the other stars. According to Vanbeveren and Conti (1980), approximately 40 percent or less of the galactic and LMC Wolf-Rayet stars show double spectra, the spectrum of a presumed OB companion being prominent in the region 3100 to 7000 Å and the emission lines looking somewhat drowned.

The classification scheme developed at Victoria by Beals (1938) separates the Wolf-Rayet stars into two sequences, which he considered to be more or less parallel insofar as effective temperature is concerned. Those stars which show dominant emission lines from H, He I, He II, N III, N IV, and N V are put into classes WN5, WN6, and WN7, while those whose spectra are dominated by lines of H, He I, He II, C II, C III, C IV, O III, O IV, and O V are put in the classes WC6, WC7, and WC8. Among the WC stars, the width of the emission lines correlates with the spectral type, the lines being broad in type WC6 spectra and relatively narrow in type WC8. In 1938, considerable emphasis was placed on the idea that the carbon and oxygen abundances were very deficient in WN stars, whereas nitrogen was significantly underabundant in WC stars. Later work has shown that, although there are wide differences in the *relative* intensities of lines from the various C, N, and O ions in Wolf-Rayet spectra of the two sequences, lines of C IV are quite strong in WN spectra and weak lines from some of the O ions also occur there, while in WC spectra weak lines from the N ions probably occur. These aspects which are questioned, together with the behavior of the ultraviolet spectrum, will be discussed later in Parts II and III. Most of the emission "lines" seen in Wolf-Rayet spectra are blends of several components. Emission in the Balmer series of hydrogen seems to be weak or absent in many Wolf-Rayet stars, but emission from the Pickering series of He II is present in all. The weakness of the Balmer lines of hydrogen can be demonstrated by noting the progression in intensity along the Pickering series. The He II lines described by the quantum numbers, 4 - n, where n is an even number, have greater intensities than those for the case in which n is odd in those stars for which emission from hydrogen is strong. However, this conclusion is based on assuming that there is no photospheric absorption line at the positions of the Balmer lines. Evidence that hydrogen absorption lines formed in the photosphere may exist has been presented by Underhill (1980). Wolf-Rayet stars have an infrared excess which is attributable chiefly to free-free emission in the neighborhood of H^+ and He^{++} ions (Hackwell et al., 1974; Cohen et al., 1975).

The spectral classification system for Wolf-Rayet stars has been revised by Hiltner and Schild (1966) on the basis of well-widened prismatic spectra covering the range from 3200 to 6700 Å. Hiltner and Schild have published spectra and given classifications for 24 WN stars, 19 WC stars, 2 miscellaneous emission-line objects, and 4 Of stars. They recognize two sequences among the WN stars: WN-A and WN-B. The WN-A stars have relatively narrow emission lines and a strong continuum, and most exhibit absorption lines characteristic of O or B stars. Many are known to be double-lined binaries. The WN-B stars have broad emission lines, and only one of them, HD 193928, is known to be a binary. The stars of the WN-B grouping show a smaller range of ionization than those of the WN-A group, but the sample is small and this difference may not be significant. It is difficult to establish useful line ratios for classifying WN stars since there are few prominent emission lines to choose from. Lines of N V tend to be weak or blended with N III; many of the He II lines blend with N III lines, and He I 5876, which is an important line in the Beals' (1938) classification system, is in absorption in WN4-A and WN5-A but in emission in the WN-B series. Hiltner and Schild mistakenly attribute the C IV blend at 5806 Å to N IV, a misidentification that originated with the Beals' classification system. They list criteria defining types WN4, WN5, WN6, WN7, and WN8.

The classification of WN stars, using as criteria only the apparent strengths of N III 4640-42, N IV 4058, and N V 4604, has been reviewed also by Walborn (1974). He noted that some WN stars show essentially identical *relative* line strengths (thus implying a similar level of excitation), but different *widths* for their emission lines. He expanded the A and B characterization of Hiltner and Schild (1966). The meaning of the widths of the emission lines in Wolf-Rayet spectra is discussed further in Part III.

The criteria used by Hiltner and Schild for defining the WC types depend chiefly on the relative intensities of lines of C II, C III, and C IV. In addition, the line at 5876 Å, attributed to He I and C III, is compared to the He II line at 5411 Å. The correlation between width of the emission lines and ionization, noted by Beals to be valid for WC stars, is confirmed. Hiltner and Schild recognize types WC5, WC6, WC7, and WC8. They note that some stars with prominent emission lines from the C ions do not fit into the WC sequence which they have defined.

A thorough review of the spectral classification of Wolf-Rayet stars from slit spectra and from narrowband photometry has been made by L. F. Smith (1968a, 1968b). Her criteria are given in Tables 1-4 and 1-5 for the WN and WC spectral types. The WN criteria depend predominantly on the state of ionization of nitrogen, and they make use of lines visible on blue-sensitive photographic emulsions. Reference is made to lines of He I only in order to differentiate between types WN7 and WN8. No use is made of the strengths of the He II emission lines. In the case of the WC stars, the line ratios, C III 5696/O V 5592 and C III

Table 1-4
Classification Criteria for WN Spectra (Smith, 1968a)*

Spectral Type	Characteristic	Criteria
WN8	N III >> N IV	He I strong with violet absorption edges; N III 4640 ≈ He II 4686; N III 5314 present
WN7	N III >> N IV	He I weak; N III 4640 < He II 4686
WN6	N III ≈ N IV	N V present but weak; N III 4640 blend present
WN5	N III ≈ N IV ≈ N V	N III 4640 blend present
WN4.5	N IV > N V	N III very weak or absent
WN4	N IV ≈ N V	N III very weak or absent
WN3	N IV << N V	N III absent

*Lines used to represent each ion: N III 4634,40,41 blend, 5314; N IV 3479-84 blend, 4058; N V 4603, 4619, 4933-44 blend.

Table 1-5
Classification Criteria for WC Spectra (Smith, 1968a)

Spectral Type	C III 5696/O V 5592	C III 5696/C IV 5806	Width of C III, IV 4650 (Å)
WC5	<1.0	0.3	85
WC6	>1.0	0.3	45
WC7	8.0	0.7	35
WC8	–	1.0	–
WC9	–	3.0	10

5696/C IV 5806, are used together with the width of the C III, C IV blend at 4650 Å. The O V and C IV lines are strongest relative to C III in type WC5 and weakest, or nonexistent, in WC9. Smith finds that stars in Beals' WC6 class can be separated into two classes: WC5 and WC6. She expands the classification scheme for the low-ionization WC stars, calling stars with spectra intermediate between Beals' classes WC7 and WC8 type WC8 and renaming Beals' type WC8 as WC9. Smith's criteria for the WC spectral types exclude reference to the strengths of lines of He I or He II.

A unique relationship, given by Smith (1968a), exists between the spectral types assigned by Hiltner and Schild (1966) and those by Smith. The differences are due chiefly to the philosophy adopted concerning how explicitly spectra showing the characteristic features of Wolf-Rayet and of O or B stars are to be recognized as those of binaries. Smith implies a binary origin for such spectra. A few stars are noted by Smith as having both WC and WN characteristics. The interpretation of these spectra will require study at as high a spectral resolution as possible and over as long a wavelength range as possible in order to unravel what is occurring. Smith has clarified and systematized greatly the spectral classification of the somewhat heterogeneous group of stars known as Wolf-Rayet stars, but some peculiar and unique objects are still included in this general category of emission-line objects.

More recently, in *The Sixth Catalogue of Galactic Wolf-Rayet Stars*, van der Hucht et al. (1981) have tabulated their classification scheme, which is based on the Smith system. New subtypes are defined, the classification criteria of which are given in their Table VII:

- WN9: N III present; N IV weak or absent; He I, lower Balmer series P Cyg.
- WN2: N V weak or absent; strong He II.
- WC8.5: C III > C IV; C II not present.
- WC4: C IV strong, C III weak or absent; O V moderate.

The subtype WN2.5 in the same scheme has been introduced by Conti et al. (1983b): N V present, N IV absent.

G. Ultraviolet Spectral Region

For more than a decade, several astronomical satellites have contributed to the data base of ultraviolet observations. We shall review briefly the main experiments as far as they concern the energy distribution and the low-resolution spectrum of O-type stars. Most of them have been used to undertake ultraviolet spectral classification. The detailed observation of the ultraviolet region in which strong resonance lines are found has shown that all O and WR stars are losing mass via a stellar wind. This aspect, as well as the general study at high resolution of the ultraviolet spectrum, will be considered later in Parts II and III of this book.

1. Main Ultraviolet Programs. The presentation of each experiment is followed by references which provide a description of the system, as well as information about the data reduction procedure and the calibration in absolute energies which cannot be given here. See also the compilation made by Underhill in *B Stars With and Without Emission Lines* (Underhill and Doazan, 1982).

The Second Orbiting Astronomical Observatory Satellite (OAO-2) was launched by the National Aeronautics and Space Administration (NASA) in 1968. The photometric experiment contained 11 medium-band photometer/filter combinations in the λλ 1330 to 4250 Å wavelength range with FWHM between 180 and 850 Å (Code et al., 1980).

The spectrophotometric experiment consisted of two medium resolution spectrum scanners operating from 1160 to 1850 Å in steps of about 10 Å, the FWHM of the instrumental profile being about 12 Å, and from 1850 to

3600 Å in steps of about 20 Å, the FWHM of the instrumental profile being about 22 Å (Code et al., 1970; Savage and Jenkins, 1972). Two atlases are available which give stellar spectra for 164 bright stars of diverse types, including 16 O stars and 2 WR stars, observed from 1200 to 3600 Å and for another 166 stars, including 10 O stars and 3 WR stars, observed either in the 1200 to 1850 Å wavelength region or in the 1800 to 3600 Å region (Code and Meade, 1979; Meade and Code, 1980). The OAO-2 photometry for 531 stars of diverse types, 50 of which are O stars and 6 of which are WR stars, is given in Code et al. (1980).

The Copernicus Satellite (OAO-3) was launched in 1972 and contained a telescope/spectrometer designed for high-resolution studies of stellar spectra. Photons were detected by four phototubes: U_1 and U_2 tubes, which are sensitive in the far-ultraviolet (second-order data) and operate in the ranges λλ710 to 1500 Å and λλ750 to 1645 Å with resolutions of 0.05 and 0.2 Å, respectively, and V_1 and V_2 photomultipliers, which are restricted to the near-ultraviolet (first-order data) and operate in the ranges λλ1640 to 3185 Å and λλ1480 to 3275 Å with resolutions of 0.1 and 0.4 Å, respectively. Because of high noise in the near-UV detectors, most of the useful data are contained in the shorter wavelength scans (Rogerson et al., 1973). A catalog of 0.2 Å resolution far-UV stellar spectra between 1000 and 1450 Å measured with Copernicus has been presented by Snow and Jenkins (1977) for 17 O-type stars and 43 B-type stars. In addition, the spectra of four WR stars are given in Johnson (1978).

The TD1 Satellite, launched by the European Space Research Organization (ESRO) in 1972, carried two experiments for observing stellar spectra: S2/68 and S59.

The S2/68 low-resolution scanning spectrophotometer had a photometric channel centered at 2740 Å (FWHM = 310 Å) and three scanning spectrometers covering the 1350 to 2550 Å wavelength region with an effective passband of width 35 to 40 Å (Boksenberg et al., 1973). The *Ultraviolet Bright Star Spectrophotmetric Catalogue* (Jamar et al., 1976) and its supplement (Macau-Hercot et al., 1978) give the observed energy distributions for 1791 stars, including 53 O-type stars and 3 WR stars. A catalog of ultraviolet energies for 31215 stars, including 97 O stars, has been published by Thompson et al. (1978).

The S59 stellar spectrophotometer observed the stars in three wavelength regions of about 100 Å wide and centered around 2110 Å, 2540 Å, and 2820 Å, with an average resolution of 1.8 Å (de Jager et al., 1974). A total sample of 220 bright early-type stars, 9 of which are O stars and 1 is a WR star, has been measured, and the corresponding atlas is given by Lamers et al. (1981).

In the summer of 1973, the Skylab S-019 Experiment utilized a 15-cm-aperture objective-prism camera to obtain spectra in the 1250 to 3000 Å region, with a resolution of about 2 Å at 1400 Å and about 12 Å at 2000 Å, for several fields containing O and B stars (Henize et al., 1975; O'Callaghan et al., 1977). Statistical data over the entire O–B2 range (133 stars, 33 of which have a type O) have been presented by Henize et al. (1981).

The Astronomical Netherlands Satellite (ANS) was launched in 1974. The ultraviolet photometer aboard covered the 1550 to 3300 Å wavelength region by means of six intermediate bands with FWHM between 50 and 200 Å (van Duinen et al., 1975; Wesselius et al., 1980a). *A Catalog of ANS Observations* containing results for about 4000 stars was published by Wesselius et al. (1980b).

The International Ultraviolet Explorer (IUE) was launched on January 26, 1978. This project has been a joint undertaking of NASA, the U.K. Science Research Council (SRC), and the European Space Agency (ESA). Two echelle spectrographs, coupled with television cameras, cover the spectral ranges 1150 to 1950 Å and 1900 to 3200 Å, respectively. Two resolution modes are available: a high-dispersion resolution of the order of 0.2 Å and a low-dispersion resolution of about 6.5 Å (Boggess et al., 1978a, 1978b). *The IUE Ultraviolet Spectral Atlas* (Wu et al., 1983) presents spectra in the low-resolution mode for both wavelength ranges of about 175 stars, including 18 O stars from O3 to O9.5. Part I of the *IUE Low-Dispersion Spectra Reference Atlas* (Heck et al., 1984) gives spectra for 229 normal stars, including 42 O stars; the short- and long-wavelength spectra have been combined with 2 Å stepping. These sets of spectra are intended to provide reference sequences of ultraviolet low-dispersion spectra with a reasonably good coverage of the Hertzsprung-Russell diagram. They are calibrated in absolute energy units with the system of Bohlin and Holm (1980). No correction for interstellar reddening has been applied.

2. Classifications in the Ultraviolet Region. Thanks to the various satellites, a large number of ultraviolet spectra are now available, and it is possible to test ultraviolet criteria that are useful for spectral classification. It may be interesting also to go further and to explore how spectral classification would look if the ultraviolet were the only spectral range known to us. A complete identity between classifications made in the visible and in the ultraviolet regions is not to be expected because the spectral lines which can be seen in these two regions differ somewhat by both their nature and their origin. It is in the ultraviolet spectral range that the strongest lines of most elements are found, and the ultraviolet classifications will probably be more influenced by anomalies in the chemical composition of the star. The lines observed in the visible range are generally formed in the photosphere. In the ultraviolet regions, strong lines originate in the outer layers of the star beyond the photosphere, and ultraviolet spectral classifications may therefore give information about nonphotospheric phenomena that are ignored by the visible classifications. In the present state of the theory, however, the nonphotospheric phenomena are not correlated with the basic parameters of the photosphere, and the most important properties of the star, such as the luminosity, the mass, and the radius, are better represented in the spectral classifications made in the visible spectral range from lines formed in the photosphere.

We must note also that the ultraviolet spectra are generally crowded with blended stellar and interstellar lines in such a way that classification spectra should have a resolution comparable to (or even better than) that of the MK system. At the time of writing, the ultraviolet spectra which are used for classification have a resolving power, $\lambda/\Delta\lambda$ (see Table 1-6), much lower than in the case of the MK system, for which $\lambda/\Delta\lambda \simeq 2000$.

Another difficulty for ultraviolet spectral classification is the incompleteness of the presently available atomic data, with the result that the percentage of unidentified lines may be as high as 50 percent. However, several attempts to classify stars from ultraviolet criteria have been made, and we briefly describe some of them which concern hot stars, among others. More recent results are also discussed by P. S. Conti in Part II of this book.

At low resolution, the spectral region longward of 1900 Å is more or less featureless for O stars or presents features which have both stellar and interstellar contributors. Inversely, in the spectral region from 1150 to 1900 Å some characteristic stellar features which are generally a blend of several lines are seen.

The mid-ultraviolet spectrum of early-type stars has been best investigated for classification purposes from the S-59 spectra which cover the 2060 to 2160, 2490 to 2590, and 2770 to 2870 Å wavelength ranges with an average spectral resolution of 1.8 Å. Three features have been selected: an absorption feature near 2078 Å attributed mainly to Fe III, an absorption feature near 2549 Å which is mainly due to Fe II, and the four Mg II lines near 2800 Å (de Jager et al., 1975; Lamers et al., 1979). Instead of measuring equivalent widths, three ratios, relative to the fluxes in a narrow band at the wavelength of the line and in a neighboring region in which the pseudo continuum is reached, have been defined: <Fe II>, <Fe III>, and <Mg II>. The results for 138 early-type stars, including 9 O stars plus γ^2 Vel

Table 1-6
Resolving Power, $\lambda/\Delta\lambda$, for Different Ultraviolet Classifications

Experiment	$\lambda/\Delta\lambda$		References
	At 1500 Å	At 2500 Å	
OAO - 2	125	115	Underhill et al. (1972) Panek and Savage (1976)
TD1 + S2/68	40	70	Cucchiaro et al. (1976)
TD1 + S59		1400	de Jager et al. (1975) Lamers et al. (1979)
Skylab S-019	~500	~900	Henize et al. (1975) Henize et al. (1981)
IUE low dispersion	250	400	Classification Program under way Heck et al. (1984) *IUE Spectral Atlas,* Wu et al. (1983)
IUE high dispersion	7500	12500	Classification programs possible*

*Note added in proof: Since this review was written, the "IUE Atlas of O-Type Spectra from 1200 to 1300 Å" by Walborn et al. (1985) has been published.

(WC8 + O), observed with a good signal-to-noise ratio lead to the conclusion that the above-mentioned parameters are not efficient for O stars: <Fe II> is a very sensitive luminosity parameter for stars later than B2, <Fe III> is a sensitive luminosity parameter for B2 to B5 stars, and <Fe III>/<Fe II> is a very sensitive temperature indicator for stars of type B2 and later (Lamers et al., 1979). Moreover, <Mg II>, which shows a strong correlation with T_{eff} for stars of type B3 and later, is severely affected by interstellar components which cannot be neglected for stars hotter than B5 (Lamers and Snijders, 1975).

From the information contained in TD1-S2/68 scans (1350 to 2550 Å), Cucchiaro et al. (1976, 1977, 1978a, 1978b) have investigated the possibility of assigning spectral types on the MK system. Their study extends from early B-type stars to late-type stars. It does not concern mainly O-type stars, but could be applied to them. For early B-type stars, Cucchiaro et al. use the ratio $r_1/r_2 = F_\lambda \simeq 1410/F_\lambda \simeq 1550$ as temperature indicator, and the ratio $r_3/r_2 = F_\lambda \simeq 1620/F_\lambda \simeq 1550$ as luminosity indicator. The feature at 1410 Å is the Si IV doublet, the feature at 1550 Å is a blend of the C IV doublet with Fe III and Si II, and the feature at 1620 Å is a blend of Fe II and Fe III.

From the OAO-2 short-wavelength spectra, Panek and Savage (1976) have identified the principal contributors to the strongest features present in the spectra of O- and B-type stars (see Table 1-7). They have also performed an extensive analysis of the behavior of the C IV $\lambda\lambda$1548, 1550 and Si IV $\lambda\lambda$1393, 1402 resonance doublets which are generally very strong in the hot luminous stars. They often show P-Cygni profiles, which are detected only in the most extreme cases at the low resolution of OAO-2. In dwarfs, the absolute strength of C IV and Si IV blends correlates well with spectral type. The C IV absorption increases sharply from B2

Table 1-7
Strongest Features Observed in OAO-2 Far-Ultraviolet Spectra of O and B Stars
(from Panek and Savage, 1976)

λ(Å)	Principal Contributors	Comments
1175	C III (1175–1176, UV4)	Maximum strength near B1, insensitive to luminosity.
1215	H I (1216, Lα) Si III (1207, UV2) N V (1239–1243, UV)	For stars hotter than B2, feature is mostly due to interstellar absorption. For cooler stars, the stellar line dominates.
1300	Si III (1295–1303, UV4) Si II (1304–1309, UV3)	Si III dominates in early B stars; Si II dominates in late B stars. The feature increases in strength toward cooler spectral types and increases slightly in strength with increasing luminosity.
1400	Si IV (1394–1403, UV1)	Maximum strength near B1; very sensitive to luminosity. P-Cygni profiles are apparent for the very luminous O stars.
1550	C IV (1548–1551, UV1)	Maximum strength occurs for stars earlier than O9; very sensitive to luminosity. P-Cygni profiles are apparent for very luminous O stars.
1600–1640	Fe III (1601–1611, UV 118) Al III (1600–1612) N II (1627–1630) He II (1640, UV12) and unidentified lines at 1621 and 1632	The strength of this broad blend of lines is relatively insensitive to spectral type but is sensitive to luminosity.
1720	N IV (1719, UV7) C II (1720–1722, UV14.02) Al II (1719–1725, UV6)	The strength of this feature is relatively insensitive to spectral type but is sensitive to luminosity (see Underhill et al., 1972).

to O9; for dwarfs earlier than O9, the observational material is too limited to infer a systematic behavior of C IV. The Si IV feature smoothly increases in strength from not measurable around B3 to a broad maximum near B1, and then decreases with increasing temperature for O stars. The equivalent width of C IV and Si IV increases by a factor of 2 or 3 when moving from dwarfs to supergiants at a given spectral type (Figure 1-3).

These general variations in the strengths of the resonance lines of C IV and Si IV have also been recognized by Henize et al. (1975) using data derived from the Skylab S-019 experiment. A classification system based on the spectra of 133 stars over the entire O–B2 range has been presented by Henize et al. (1981). They find the same qualitative behavior of Si IV and C IV absorption strengths with temperature and luminosity as observed by Panek and Savage (1976). Henize et al. also consider the appearance of P-Cygni profiles in the hottest and most luminous stars. They show that C IV emission occurs when the type is earlier than O7–O8 for dwarfs, O9–O9.5 for giants, and B1–B2 for supergiants. As a matter of fact, the

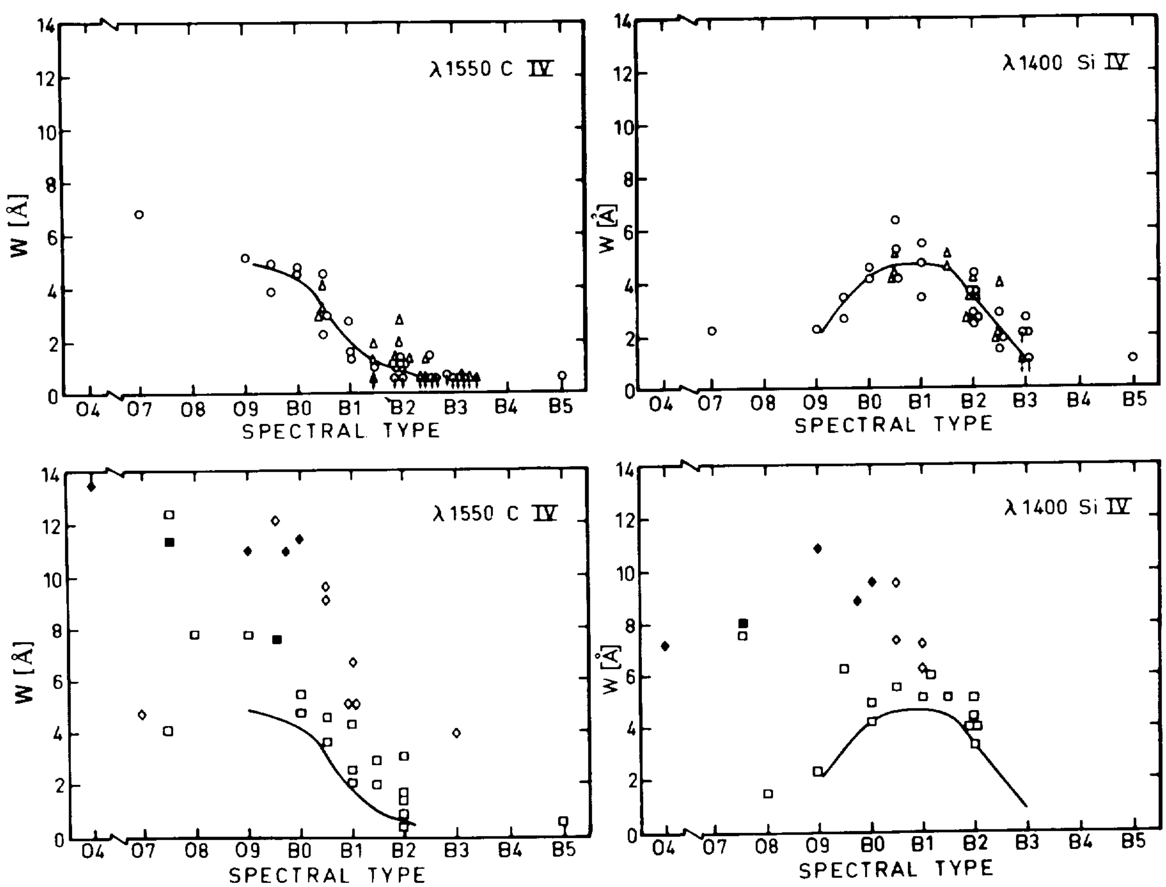


Figure 1-3. The relation between the equivalent width of the C IV (1550 Å) and Si IV (1400 Å) absorption features and the stellar spectral type (adapted from Panek and Savage, 1976). The measured values for normal stars are plotted individually. (Be and Oe stars, Bp stars, stars with known variability of ultraviolet line strengths, and known β Cephei variable stars are not plotted.) The different symbol shapes denote different luminosity classes according to the code: O — luminosity class V; Δ — class IV; □ — class III or II; ◊ — class I. The arrows on the data points denote upper limits. Values corresponding to lines showing emission at the OAO-2 resolution are indicated with a solid symbol.

C-2

C IV emission turn-on takes place at a given C IV absorption intensity—5 on their scale (see Figure 1-4). The Si IV emission is restricted to O4–B0 supergiants. Figure 1-5 shows that O stars with no Si IV absorption must be classified as O V, stars with weak to moderate Si IV absorption (and no Si IV emission) are of luminosity class III, and stars with strong Si IV absorption together with Si IV emission are clearly to be classified as supergiants. In conclusion, the authors propose the use of the log (Si IV/C IV) versus C IV diagram as the primary classification tool (Figure 1-6). Additional criteria (e.g., the presence of C IV or Si IV emission) may then be invoked to refine this classification.

The log (Si IV/C IV) versus C IV diagram for OAO-2 data is similar to Figure 1-6 in spite of the large difference in spectral resolution between the two sets of data. This shows the usefulness of the classification criteria used. However, the precise relationship between specific areas of the log (Si IV/C IV) versus C IV diagram and MK classes must be expected to vary with the resolution of the available spectra. The presence or absence of emission depends still more on the resolution.

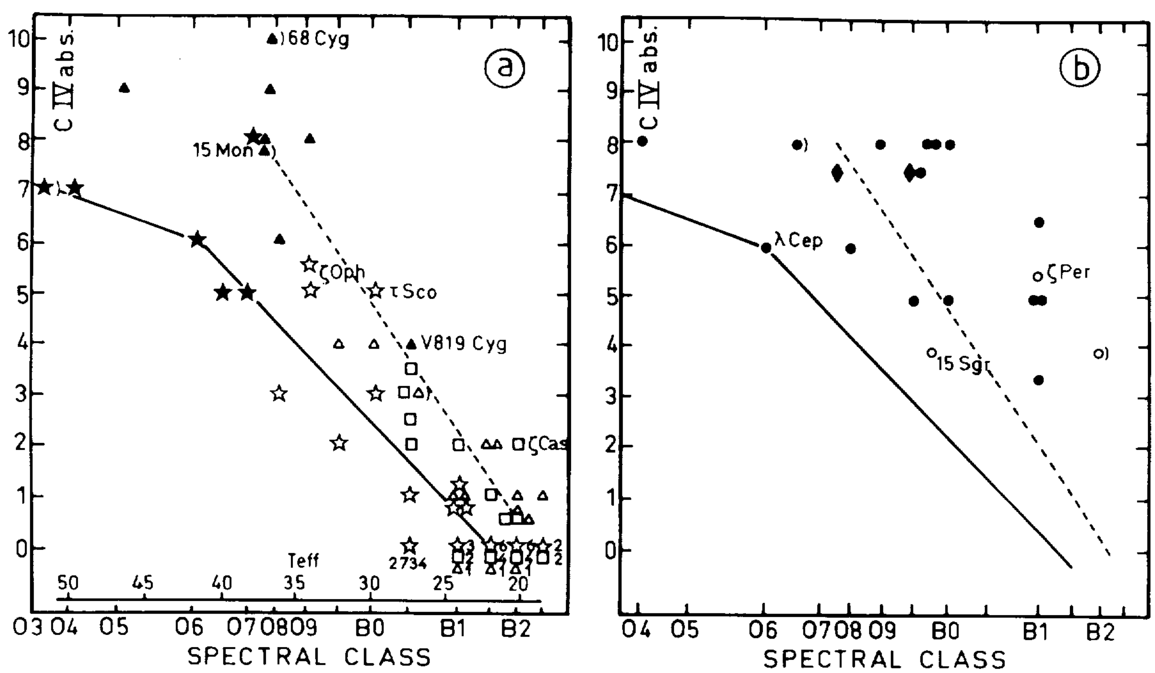


Figure 1-4. (a) The relationship between C IV absorption strength and spectral type for luminosity classes III-V (☆ — luminosity class V, □ — luminosity class IV, and Δ — luminosity class III). Closed symbols designate stars with emission at C IV. A parenthesis designates data derived from IUE spectra. The solid line is the estimated median line for class V stars; the dashed line is the estimated median line for class III stars. The scale of spectral classes is distorted to provide a linear scale of effective temperature. This scale is indicated just above the lower border. For C IV absorption strength zero at spectral class B1 and later, the number following the luminosity symbol indicates the number of such stars superposed at that point. (b) The relationship between C IV absorption strength and spectral type for luminosity classes I and II (◊ — luminosity class II and O — luminosity class I). Closed symbols and parentheses have the same meaning as in (a). The median lines for class V and III stars have been transferred from (a). (From Henize et al., 1981.)

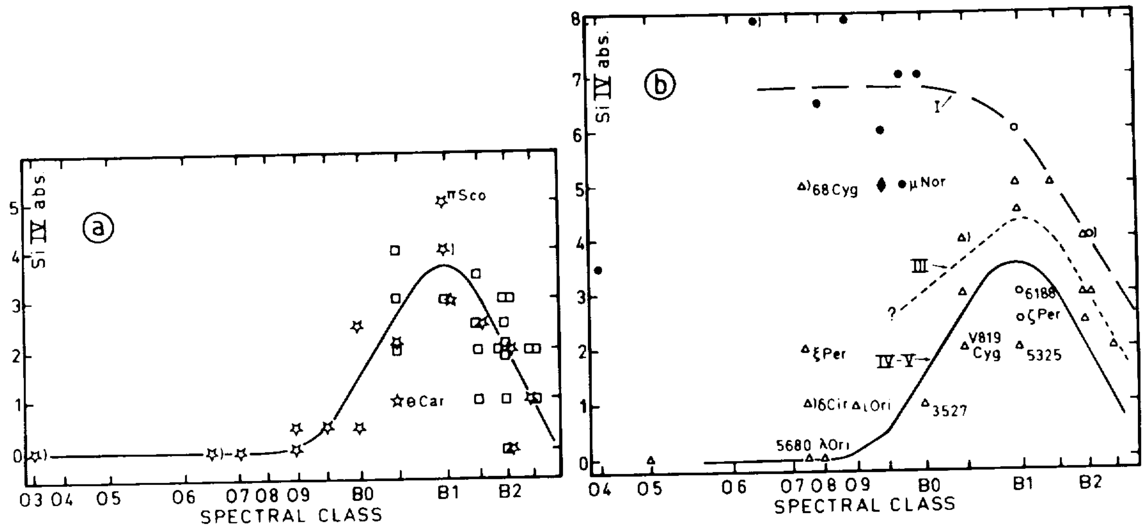


Figure 1-5. (a) The relationship between Si IV absorption strength and spectral class for luminosity class IV and V stars. The symbols are the same as those in Figure 1-4. (b) The relationship between Si IV absorption strength and spectral class for luminosity class I–III stars. The symbols are the same as those in Figure 1-4, except that closed symbols designate stars with Si IV emission. The median line from (a) has been transferred to this diagram for comparison purposes (solid line). The short-dashed line is the estimated median line for class III stars, and the long-dashed line is the estimated median line for supergiant stars. (From Henize et al., 1981.)

According to the experience of Henize et al., the low-dispersion IUE spectra fit the S-019 calibration very well.

The aim of all the classification efforts described above is to reproduce as well as possible the MK classification. Another possibility is to ignore the visible spectral range and try to establish a classification system in the ultraviolet from scratch. A first attempt in this direction was made by Cucchiaro (1982), using the material provided by the S2/68 experiment. The MK background is still very strong in this work, and the number of spectra (about 1900) is too small to establish a really independent ultraviolet classification. However, some general results were obtained, which are discussed in Jaschek and Jaschek (1982). The first point is that the ultraviolet spectra can be separated into two groups—the normal spectra and the abnormal spectra—like in the visible region. The ultraviolet normal group and the visible normal group have most of their stars in common but not all (see Figure 1-7). Another point is that there is a higher percentage of abnormal spectra in the ultraviolet than in the visible spectral range, although the ultraviolet resolving power is much smaller than the visible one. With a better resolution in the ultraviolet, the percentage of abnormal stars can only increase.

A really independent ultraviolet spectral classification can be based only on the examination of many thousands of spectra having a resolution comparable to the MK resolution. The IUE observations have provided these thousands of spectra. After the workshop on *UV Stellar Classification* held in October 1981 at VILSPA (ESA IUE satellite tracking station), a working group was created whose main task is to provide in the near future sequences of calibrated low-dispersion spectra covering all spectral types, and an ultraviolet stellar classification based on these data is described by Heck et al. (1984). The resolving power of the IUE low-dispersion spectra provide a considerable improvement over the S2/68 data but is still lower than the resolving power of the MK spectra, and Baschek (1982) suggested that IUE high-resolution spectra be used for ultraviolet stellar classification (see note to Table 1-6 and Part II).

III. INTRINSIC COLORS, CIRCUMSTELLAR REDDENING, AND INTERSTELLAR REDDENING

Since most O and Wolf-Rayet stars lie in the plane of the Galaxy, the effects of wavelength-dependent interstellar extinction on their spectra must be taken into account. We shall first discuss observations made at the surface of the Earth. In the last section, we give information on the intrinsic colors of stars in the ultraviolet region.

A. The Problem of the Intrinsic Colors of the O Stars

It is a basic assumption underlying the use of spectral types and intrinsic colors for stars (e.g., in statistical studies) that the shape of the radiation field from a star may be specified uniquely by means of the spectral type of the star and *vice versa*. Thus, one expects to be able to establish a single-valued relationship between spectral type and intrinsic color. This has proved to be possible for stars in most parts of

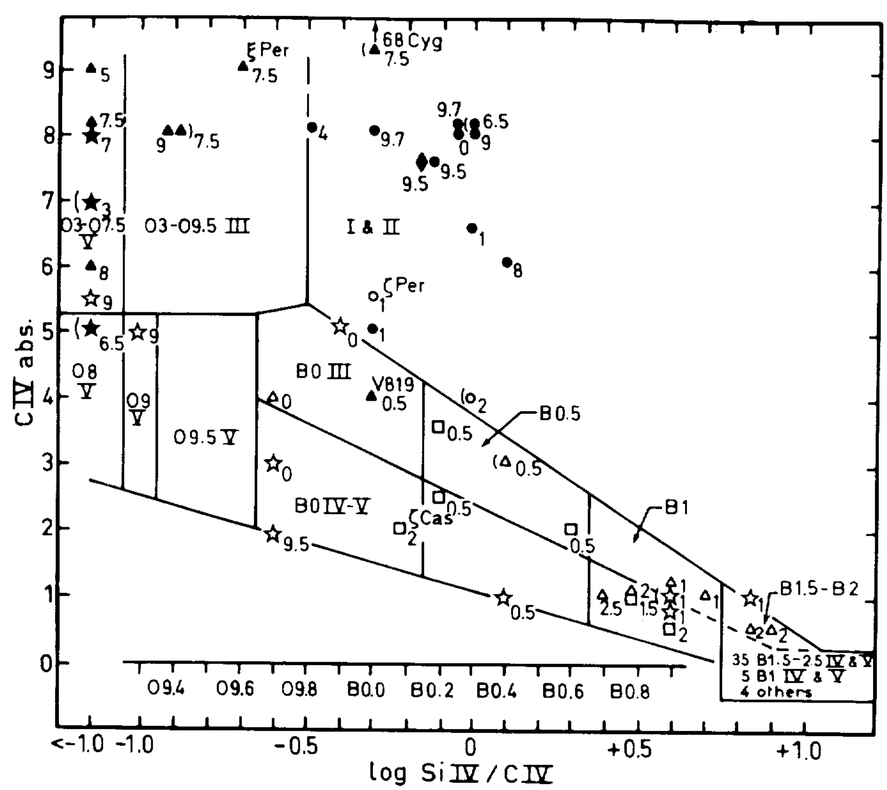


Figure 1-6. The log (Si IV/C IV) versus C IV diagram for S-019 data. The symbols are the same as those for Figure 1-4. Closed symbols denote stars with C IV emission. The number near each symbol gives the temperature class. Lines divide the diagram into areas corresponding to MK classifications. Boundaries where the separation is weak or uncertain are shown by dashed lines. The spectral types of 44 stars with zero C IV strengths are described in the box at the lower right. (From Henize et al., 1981.)

the HR diagram. However, the results obtained for the O stars have been ambiguous, and several questions arise, including the very basic question of whether spectral classifications based on observations made in the limited spectral region between 4000 and 4800 Å can indeed be uniquely and monotonically related to the total energy field from the star (i.e., to the effective temperature of the star). In particular, one realizes from various spectroscopic details that a significant number of O stars may be shrouded by extended atmospheres having a variety of physical characteristics and that these extended atmospheres may modify the shape of the continuous spectrum observed in the spectral region accessible from the surface of the Earth. On the other hand, since well-defined relationships between spectral type and intrinsic color exist for normal stars of spectral types B, A, and later, the problem of determining such relations for O stars presents real interest. In the following, we shall summarize what has been achieved to date.

In the case of the O stars, even the closest stars are sufficiently distant that one cannot ignore the effects of interstellar extinction. The observed spectral distribution of the stellar energy—whatever method is used, be it broadband photometry, narrowband photometry, or spectrophotometry at high or low resolution—is affected by this extinction and does not directly give information on the intrinsic colors of the O stars. In addition, the possibility of the existence of a reddening, called "circumstellar," caused by an envelope of dust or gas which may be inferred to be present because of an abnormal energy distribution in the infrared in the case of some O stars, or because of an abnormal energy distribution in the region of the Balmer discontinuity in some emission-line O stars (e.g., HD 60848 and HD 45314) complicates the situation even further.

For a long time, these envelopes, whose presence was not recognized, and the rather imprecise spectral classifications that were available for O stars have falsified the laws of interstellar extinction obtained by photometric methods, with the result that many have be-

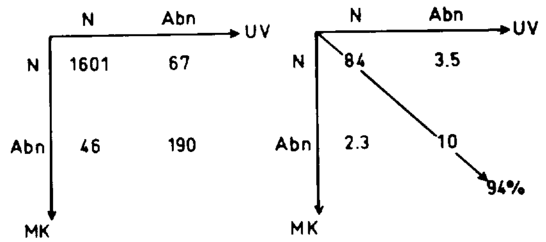


Figure 1-7. Number and percentage of normal and abnormal stars in UV and MK classifications of a sample of 1904 stars. (From Jaschek and Jaschek, 1982.)

lieved that important variations in the character of the interstellar extinction law occur from one region of the Galaxy to another. This conclusion has been reached despite the contrary results found earlier by Stebbins and Whitford (1945) and by Divan (1954, 1956, 1963). Divan took great care in her choice of stars to eliminate discordant factors like circumstellar shells.

The question of the uniformity of the interstellar extinction law has been rediscussed by Underhill (1966), and her conclusion that, for the wavelength range extending from the infrared to the near ultraviolet ($\lambda > 3000$ Å), the law of interstellar extinction is universal throughout our Galaxy has generally been confirmed by the observations made since that review (see for example, Hackwell and Gehrz, 1974; Wiemer, 1974; Schultz and Wiemer, 1975; and Nandy et al., 1975, 1976). Except for very special regions, such as the Orion nebula, dense molecular clouds, or circumstellar envelopes, which are not representative of the general diffuse medium, the interstellar extinction law seems to be the same everywhere for wavelengths longer than about 3000 Å. Variable extinction laws in the far-ultraviolet ($\lambda < 3000$ Å) are discussed by P. S. Conti in Part II of this book; they introduce an uncertainty in the ultraviolet intrinsic colors but have no influence on the optical regions which are discussed below.

The fact that, to a good approximation, interstellar extinction is uniform simplifies considerably the problem of obtaining intrinsic colors. In this case, well-defined and almost linear color relations exist in each photometric system for stars of a given spectral type, and from these relations, if the intrinsic color is known for one particular index, the intrinsic colors for the other ones can be deduced directly. The problem of obtaining intrinsic colors for the O stars still remains difficult. Each author has chosen to use one or another of the available methods according to his personal preference and to the possibilities of his instrumentation.

B. Principal Methods for Determining Intrinsic Colors

Here we shall describe the three chief methods which are used to find intrinsic colors for O stars in the visible and infrared wavelengths.

1. Extrapolation of Results Obtained for B Stars. In the case of the main-sequence B stars, one can show that the bluest stars in each spectral type suffer negligible interstellar extinction and that they therefore permit one to define a relation between intrinsic color and spectral type. Some typical examples of color-spectral-type diagrams are shown in Figures 1-8 and 1-11. In the case of the B stars, the limit to the distribution of points formed by the bluest stars traces a relatively regular curve that corresponds to temperatures which become higher as the spectral type becomes earlier, a type of variation which is suggested by the changes in the line spectrum. These results indicate that the bluest B stars are probably almost unreddened.

In the case of the O stars, even the bluest of them have observed colors that are redder than the colors of unreddened B0 stars. This is true in all photometric systems and for spectrophotometry. Making the hypothesis that this result is due to the pervasive action of interstellar extinction which affects even the most nearby of the O stars, one extrapolates the intrinsic color-spectral-type relation of the B stars, arbitrarily shaping the relation to conform to the way in which the colors of a blackbody would vary as its temperature increased from the temperatures appropriate for B stars to those believed to be appropriate for O stars. The solutions adopted by Morgan et al. (1953), by Johnson (1966), and by Schmidt-Kaler (1965) are shown in Figure 1-8. Because these extrapolations are arbitrary, attempts have been made to find the intrinsic colors of O stars by other methods.

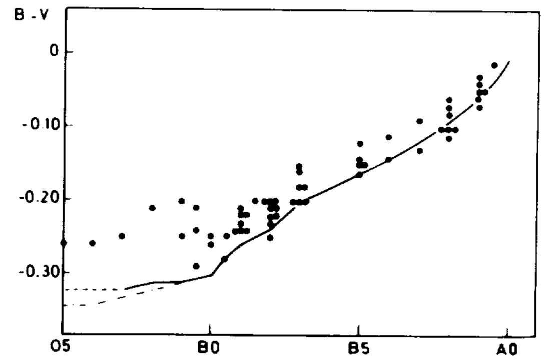


Figure 1-8. The bluest stars in the B−V color system shown as a function of the MK spectral type. The $(B-V)_0$ versus spectral-type relation adopted by Morgan et al. (1953) is shown by a full line, that of Johnson (1966) by a dashed line, and that of Schmidt-Kaler (1965) by a line of dots and dashes.

2. Method of Clusters. The O stars are often grouped in the sky in clusters which contain stars having spectral types later than O. These stars have known intrinsic colors. The distribution of the interstellar extinction in front of and within the cluster is studied by means of these stars. In relatively simple cases in which the extinction is relatively uniform over the area covered by the cluster and in which it is known which stars are members of the cluster and which are not, one can correct the observed colors of the O stars for interstellar extinction, using the information found from stars of type B and later, and obtain intrinsic colors for the O stars. However, the ideal necessary conditions are never realized, and the intrinsic colors found in this way are not so precise as one would wish them to be. Indeed, it has not been possible to improve in this way the original set of intrinsic colors for O stars established in 1953.

3. Method of Multiple Systems. Since the method of clusters has proved to be ineffective, an attempt has been made to resolve the problem by studying multiple systems containing O stars. Since the dimensions of multiple systems are much smaller than those of clusters containing O stars, one can expect that the chance that the interstellar extinction is uniform over the volume occupied by the group of stars is much greater than is the case for open clusters. A study of this sort was undertaken by Burnichon (1975), who studied a small number of multiple systems, collecting for each of them as large as possible a body of data, including photometric observations (*UBV*) and spectrophotometry in the BCD system, which gives precise information about the energy distribution on both sides of the Balmer discontinuity, together with a quantitative spectral classification of the stars. With this information in hand, Burnichon was able to reject optical systems. The intrinsic $(B-V)_0$ colors found in this study are redder than the classical (1953) values, being in the range -0.29 to -0.25 in place of -0.32 to -0.30. One faces the decision that either these are really the intrinsic colors of O stars or that these O stars are suffering a residual circumstellar reddening. Unfortunately, the number of O stars studied with this method is very small.

In practice, whatever the photometric or spectrophotometric system used, the intrinsic colors of the O stars are obtained by extrapolating from a relation that has been established for B stars which are sufficiently close to the Sun that the interstellar extinction suffered by these B stars is negligible. One defines this relation in such a fashion that it is as regular as possible; in certain cases, it may be nearly linear. However, it is important to remember that the character of the extrapolation to O stars is essentially arbitrary.

C. Intrinsic Colors of Normal O Stars in Three Photometric Systems

The number of photometric systems in use in astronomy is very large. In what follows, we shall give results for three of them, one representative of wide-band systems ($\Delta\lambda \simeq$ 1000 Å), one of intermediate-band systems ($\Delta\lambda \simeq 250$ Å), and one of very narrow-band systems (spectrophotometry with a resolution of 10 Å on the average).

1. The *UBVRIJKLMN* System. This system is composed of ten wide bands having the effective wavelengths given by Johnson (1966). The effective wavelengths (in μm) are as follows:

U	*B*	*V*	*R*	*I*
0.36	0.44	0.55	0.70	0.90
J	*K*	*L*	*M*	*N*
1.25	2.2	3.4	5.0	10.2

The transmission curves of the first three filters are shown in Figure 1-10.

The procedures used to obtain intrinsic colors in this system have been described several times (see, for instance, Johnson and Morgan, 1953; and Johnson, 1966). In the $(U-B)$ vs. $(B-V)$ diagram (Figure 1-9), the bluest B stars (which are the closest B stars) fall on a regular curve which is practically a straight line. This curve defines the intrinsic colors, $(U-B)_0$ and $(B-V)_0$, of the B stars. The bluest O stars have $(B-V)$ colors which are redder than the value of $(B-V)_0$ for stars of type B0, and if one makes the hypothesis that the bluest O stars are not reddened by interstellar extinction, one must conclude that the relationship between $(U-B)_0$ and $(B-V)_0$ ceases to be linear for spectral types earlier than B0 and that it curves to the right in Figure 1-9. However, since even the bluest O stars are relatively distant, it is reasonable to consider that this curvature is a result of interstellar extinction, and one makes the hypothesis that the linear relation between $(U-B)_0$ and $(B-V)_0$ defined by the B stars is also valid for the O stars. To find the intrinsic colors of the O stars under this hypothesis, one

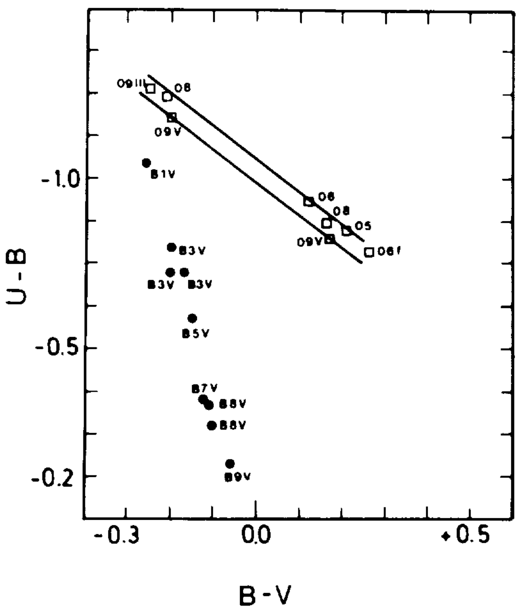


Figure 1-9. The two-color (B–V) vs. (U–B) diagram for the bluest main-sequence B1–B9 stars (filled circles) and for O stars with different amounts of reddening (squares). The reddening path for the O stars is marked. (From Johnson and Morgan, 1953.)

draws a reddening line from the observed position of an O star in the $(U\text{–}B)$ vs. $(B\text{–}V)$ diagram until it cuts the extended straight line which defines the relationship between $(U\text{–}B)_0$ and $(B\text{–}V)_0$ for B stars. The slope of the reddening line is determined by the shape of the interstellar extinction law. It can be defined only if this law is the same for all stars.

Since the work of Johnson and Morgan (1953), a large number of O stars has been observed with wideband photometry, and it has been possible to derive intrinsic $(B\text{–}V)$ for the several O subtypes by the method described above. The slopes of the reddening lines for the several color combinations have been used in finding the other intrinsic colors. The presently adopted values are generally those of Johnson (1968), which are given in Table 1-8. In the two-color diagrams of Johnson and Borgman (1963) and of Johnson (1966), the dispersion of the points is rather large, in part because of a few local variations in the shape of the interstellar extinction law and in part because of various amounts of emitting and/or absorbing material around some of the stars (see the discussion made by Wiemer, 1974; and by Schulz and Wiemer, 1975). However, the number of slightly reddened stars is generally sufficient that each intrinsic color can be determined without ambiguity if the relation between $(U\text{–}B)_0$ and $(B\text{–}V)_0$ assumed for O stars is correct and gives an unbiased relation between $(B\text{–}V)_0$ and spectral type.

FitzGerald (1970) redetermined the intrinsic colors of O stars by a method similar to that of Johnson (1968). These colors are based on a larger number of stars, but they are not entirely independent of Johnson's colors in the sense that FitzGerald adopted, as a starting point, the Johnson relation between $(B\text{–}V)_0$ and spectral type.

Barlow and Cohen (1977) have made infrared measurements for a group of O, B, and A supergiants, and they have published diagrams of $(B\text{–}V)$ vs. $(V\text{–}K)$ and $(B\text{–}V)$ vs. $(V\text{–}L)$. The dispersion of their points about a straight line is very small, even in the case of strongly reddened stars. This indicates that the wavelength dependence of the interstellar extinction is uniform in all directions in the Galaxy and that the variation with spectral type of the intrinsic colors is essentially parallel to the reddening lines. This second constraint is certainly true for the supergiants earlier than A5 which have been observed by Barlow and Cohen; it is not true for the stars of later spectral type.

Barlow and Cohen give the following linear relations between the colors:

$$V\text{–}K = 2.83(B\text{–}V) - 0.04 \quad , \quad (1\text{-}2)$$

and

$$V\text{–}L = 2.96(B\text{–}V) + 0.04 \quad . \quad (1\text{-}3)$$

These relations are valid for types A5 through O9.5. They allow us to make a comparison with the intrinsic infrared colors of Johnson (1968). Adopting $(B\text{–}V)_0 = -0.27$ for type O9.5 I (see Table 1-8), the above relations

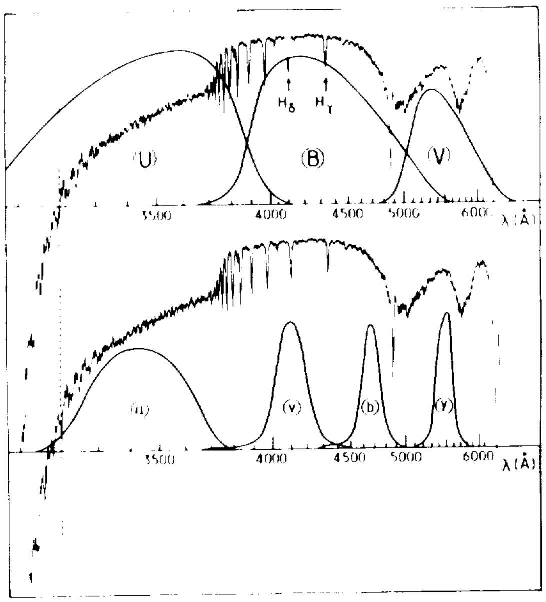


Figure 1-10. The transmission curves of the bands of the UBV and of the Stromgren uvby photometric systems, together with a microphotometer tracing of a B3 V star (ι Her). No contribution of the Paschen continuum affects the u filter. We note also that the part of the stellar spectrum which is strongly affected by the atmospheric ozone (below about 3200 Å, to the left of the shaded line) has a quite negligible contribution to the flux in the u filter, which is different from what happens in the Johnson U band. Note that the microphotometer tracing represented here is that of a photographic spectrum. The bumps observed in the green and red spectral regions are due to the sensitivity curve of the plate.

give $(V\text{-}K)_0$ = -0.80 and $(V\text{-}L)_0$ = -0.76 for type O9.5 I. The values of Johnson are -0.86 and -0.84, respectively. The differences between these results are typical of the uncertainties which still exist in the intrinsic wideband infrared colors. They may be due in part to slightly different effective wavelengths for the filters given the same names, and Koornneef has presented a series of papers which aim at enhancing the Arizona near-infrared photometric system of Johnson by homogenization of data and unification of results from various authors.

One will find in Koornneef (1983) the more recent set of intrinsic colors, effective wavelengths, and absolute calibrations relating to the J, K, L, and M bands, as well as to the H band centered at 1.65 μm, which is now in widespread usage. For O stars, the (V-infrared) indices of Koornneef are redder than those of Johnson, the differences becoming significant for late O supergiants in the K and L bands (0.1 to 0.2 mag).

2. The *uvby* β System. This photometric system, designed by Strömgren, consists of four bands of intermediate width (180 to 300 Å) distributed in the visible and near-ultraviolet part of the spectrum. The effective wavelengths of the bands are 3500 Å for u, 4100 Å for v, 4670 Å for b, and 5470 Å for y. These bands have been chosen to facilitate the quantification of the energy curves of stars of types F to O and to measure the intensities of lines in often-studied parts of the spectrum. In addition, a narrow filter centered at Hβ is used with a broad filter at Hβ to obtain an index linked to absolute magnitude. A description of the system and many of its applications has been given by Strömgren (1966).

Values of *uvby* and β have been measured for many stars of types F to O, and listings of their values can be obtained from the Centre de Données Stellaires at Strasbourg, France.

The transmission of the different filters as a function of wavelength has been given by Crawford and Barnes (1970); these transmission functions are shown in Figure 1-10, together with a microphotometer tracing of a B-type spectrum. One sees that, in contrast to what is observed for the Johnson U band, no contribution of the Paschen continuum affects the filter u. For B and O stars, the index b–y is relatively unaffected by line blanketing, and the combination of indices $c_1 = (u\text{-}v) - (v\text{-}b)$, which is relatively insensitive to interstellar extinction because the effects on u–v partially compensate for those on v–b, is a good indicator of the size of the Balmer jump. The combination $m_1 = (v\text{-}b) - (b\text{-}y)$ permits one to make a quantitative estimation of the

Table 1-8
Intrinsic Colors of O Stars According to Johnson (1968)

Spectral Type	Lum. Class	*U–V* mag	*B–V* mag	*V–R* mag	*V–I* mag	*V–J* mag	*V–K* mag	*V–L* mag
O5, O6, O7	III–V	-1.46	-0.32	-0.15	-0.43	-0.73	-0.94	-0.92
O8, O9	Ia, Ib	-1.41	-0.29	-0.15	-0.43	-0.73	-0.94	-0.92
	III–V	-1.44	-0.31	-0.15	-0.43	-0.73	-0.94	-0.92
O9.5	Ia	-1.37	-0.27	-0.13	-0.42	-0.68	-0.86	-0.84
	Ib	-1.36	-0.27	-0.13	-0.42	-0.68	-0.86	-0.84
	III–V	-1.40	-0.30	-0.14	-0.42	-0.73	-0.94	-0.92

strength of the absorption lines in the region near 4100 Å. The index m_1 is very useful for studying A stars, but is less useful for O stars.

As can be seen from Figure 1-10, filter *b* is centered on the wide emission lines which occur at 4634–41 and 4686 Å in Of stars. The effect of these emission lines on the magnitude measured at *b* explains the position of the Of stars in the c_1 versus b-y and m_1 versus b-y diagrams of Crawford (1973).

All of the indices used with the *uvby* system are affected to some degree by interstellar extinction. The problem of determining the intrinsic colors presents the same difficulties as for the *UBV* system except for the simplification that the effective wavelengths of the filters, because of the relatively narrow widths of the passbands, vary very little with increasing extinction, and the relations between the different color excesses are strictly linear.

A definitive determination of the intrinsic colors of the O stars in the *uvby* system has not yet been made. Crawford (1970, 1973, 1975a, 1975b) has given a provisionary solution which consists of extrapolating the relation,

$$(b\text{-}y)_0 = -0.116 + 0.097\, c_0 \quad , \qquad (1\text{-}4)$$

which is valid for the unreddened B stars. At type B0, c_0 is -0.06, while at type B9, it is near +0.8. In Table 1 of Crawford (1978), values of $(b\text{-}y)_0$ are given for c_0 = -0.10, -0.15, and -0.20. It is implied that these values of c_0 correspond to O stars, but no calibration in terms of spectral type has been given.

The slopes of the reddening lines in two-color diagrams derived from *uvby* photometry are easier to determine than the intrinsic colors themselves, and values have been published in several places. Values for the principal combinations of color differences used are given in Table 1-9. The best determined value is the slope of the line in the c_1 versus b-y diagram. Using this value with the relation between $(b\text{-}y)_0$ and c_0 valid for the B stars permits one to find c_0 and $(b\text{-}y)_0$ for any star if it is of type O or B and is not a supergiant. From this, one finds the color excess $E(b\text{-}y)$ which characterizes the interstellar extinction suffered by the star. The slopes $E(m_1)/E(b\text{-}y)$ and $E(u\text{-}b)/E(b\text{-}y)$ seem to be less well defined than $E(c_1)/E(b\text{-}y)$. Consequently, dereddening the indices m_1 and u-b is probably less certain than dereddening c_1.

One way of obtaining an estimate of the values of the intrinsic indices for O stars in the

Table 1-9
Relations Between the Color Excesses of the *uvby* Photometric System

Ratio	Strömgren (1966)	Crawford (1973)	Crawford (1975a, 1975b)
$E(c_1)/E(b-y)$	0.20	0.2	0.20
$E(m_1)/E(b-y)$	-0.18	-0.3	-0.32
$E(u-b)/E(b-y)$	1.84	1.6	1.5

Strömgren system would be to use calculated intrinsic colors from model atmospheres. Mihalas (1972a, 1972b) and Mihalas and Hummer (1974a, 1974b) have made high-temperature model atmospheres corresponding to O stars, and they have calculated the indices in the Strömgren photometric system for their models, LTE and non-LTE, with and without extended atmospheres. It is very difficult to conclude which model corresponds to which spectral type. One can say, however, that their LTE models, especially those with extended atmospheres, are excluded because these models give Balmer discontinuities in emission, at least for the smaller gravities, and this is not observed to be true for O stars. On the other hand, Relyea and Kurucz (1978) have calculated Strömgren indices for Kurucz' (1979) line-blanketed model atmospheres. The hot models of this series, which might correspond to O stars, do not show the Balmer jump in emission. These values for the four-color indices may give some guidance in the choice of intrinsic values for the O stars.

3. The BCD System. In this photometric system (Barbier and Chalonge, 1941; Chalonge and Divan, 1952, 1973), the distribution in the continuous spectrum of stars is measured between 3150 and 6200 Å using specially widened prismatic spectra and the techniques of photographic photometry. The spectral resolution is 10 Å on the average. In the BCD spectra, the dominant absorption and emission lines in stars of types O, B, and A are readily seen, and it is easy to define the continuum in all wavelength regions observed. The continuum is measured at about 50 regularly spaced wavelengths and is then corrected for extinction by the Earth's atmosphere. Because the resolution of 5 Å obtained in the ultraviolet allows one to determine the quantity of atmospheric ozone accurately, one can determine precise energy distributions to 3150 Å.

Relative energy distributions are determined, the comparison source being a selected standard star. The results are generally presented as a table of values of log i/i_{std} at the 50 measured wavelengths (see, for instance, Divan, 1954, 1966). However, because log i/i_{std} can be represented by a few linear functions of $1/\lambda$ over the wide spectral region studied, the energy distribution is also quite well described in a more condensed way by a few "relative gradients,"

$$G = -2.30 \frac{d}{d(1/\lambda)} \log \frac{i}{i_{\text{std}}} , \qquad (1\text{-}5)$$

which represent the slopes of these linear relations. For the O stars, two gradients suffice: G_{rb} representing the energy distribution in the 4000 to 6200 Å range, and G_{uv} for the 3150 to 3700 Å range.

To put the relative energy distribution on an *absolute* scale, stars are compared to a blackbody for which the energy distribution is given by the Planck formula, $B(\lambda, T) = c_1\lambda^{-5}$ (exp

$(c_2/\lambda T)-1)^{-1}$, with $c_2 = 14320$ microns degree. The slope of this energy distribution in the (log B, $1/\lambda$) plane is:

$$\frac{d}{d(1/\lambda)} \log B = 2.172\lambda - 0.4343\,\Phi \quad ,(1\text{-}6)$$

with

$$\Phi = \frac{c_2}{T}\left(1 - \exp(-c_2/\lambda T)\right)^{-1} \quad . (1\text{-}7)$$

The function Φ is the "absolute gradient" corresponding to the temperature T, and for two blackbodies at temperatures T_1 and T_2, the "relative gradient,"

$$G = -2.30 \frac{d}{d(1/\lambda)} (\log B_1 - \log B_2) \qquad (1\text{-}8)$$

is equal to $\Phi_1 - \Phi_2$. The function Φ has been tabulated by Kienle (1941). It varies slowly with λ, and when $\lambda T << c_2$, Φ is nearly equal to c_2/T.

To a first approximation, it is found that the O and B stars radiate like blackbodies in the separate intervals corresponding to G_{rb} and G_{uv}. Therefore, one can define absolute gradients Φ_{rb} and Φ_{uv} for the standard stars and thus fix the scale of the relative gradients.

In the first determinations (see Barbier and Chalonge, 1941), absolute gradients were determined for each program star by comparison to a hydrogen lamp calibrated against a blackbody in the laboratory of Prof. Kienle (Göttingen). But these absolute measurements are very difficult to make, and they introduce large uncertainties in the many cases in which only the much more precise relative energy distributions are needed (such as spectral classification or the determination of the shape of the interstellar reddening law). Thus, the normal procedure is to determine the energy distribution of all stars relative to a standard star and then do the absolute calibration of the standard star.

Between 1950 and 1961, the absolute calibrations were repeated several times using more precise methods than those of Barbier and Chalonge (1941). They were made first for η UMa and α Lyr, then for a set of O and early B stars (including 10 Lac, S Mon, ϵ Per, γ Ori, and γ Peg), which are more suitable than η UMa and α Lyr for a comparison with the blackbody radiation because their Balmer discontinuities are small. The terrestrial sources of light used in the stellar measurements of these O and B stars were calibrated against the blackbody of Prof. Kienle in Heidelberg by Labs and Mehltreter (Mehltreter, 1961). A great part of this material has been reduced, and values of log $I_{\text{S Mon}}$ - log B (20,000) are given in Divan (1965). The *relative* energy distributions which are published may thus be put on an absolute scale.

These new calibrations gave scales close to those adopted by Chalonge and Divan (1952). No modification of the zero point of the gradients of 1952 is therefore necessary, and all the published values of Φ_{rb} and Φ_{uv} are on the same scale. *Absolute* gradients have been given for 29 well-observed stars by Divan (1966). Hayes (1970) has shown that these values agree very well with his own absolute calibrations.

The Balmer discontinuities are always measured in a differential fashion, relative to a standard star. The zero point of the scale of D has been fixed by a special procedure, independent of any absolute calibration, and described in Chalonge and Divan (1952). This method gives a result which is more precise than the absolute calibration of the standard star. The zero point has been revised since 1952, but the resulting correction of +0.012 dex has no influence for applications to spectral classification and it has not been used. On the other hand, if one wishes to compare measured discontinuities with those calculated for models, it is necessary to make this correction and to increase all the published values of D by +0.012 dex.

In the BCD system, the intrinsic colors of the O stars have been obtained by means of the extrapolation shown in Figure 1-11 (taken from Chalonge and Divan, 1973). The variable s (abscissa) is a continuous parameter for spectral type defined by means of the parameters λ_1 and D. In the diagram of Φ_{rb} against s, the bluest B stars lie on a very smooth curve. A possible extrapolation of the curve is indicated

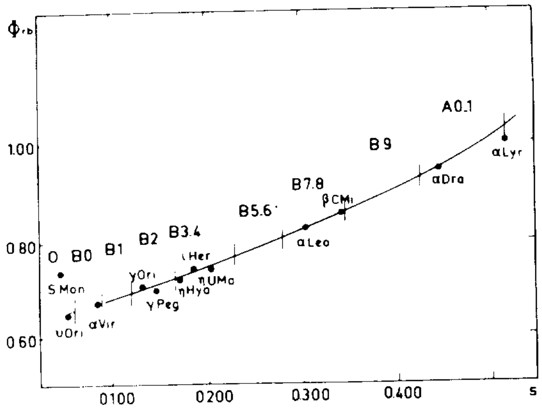


Figure 1-11. The bluest stars in the BCD system. The gradient Φ_{rb} for these stars is a very regular function (full line) of the continuous parameter of spectral type, s, defined by Chalonge and Divan (1973). The dotted line indicates a possible extrapolation for the O stars, although all O stars are situated well above this curve (see text).

by the broken line. The position of the point corresponding to υ Ori is uncertain, owing to the large zenith distance of this star. The bluest O stars lie above the projected curve. This may be the result of interstellar reddening, but a study of multiple systems (Burnichon, 1975) has shown that it is possible that the intrinsic colors of certain O stars are redder than the projected curve indicates. The value of the intrinsic gradient $\Phi_{\mu v}$ can be deduced from that of Φ_{rb} and knowledge of the shape of the interstellar extinction law.

The relation between *s* and spectral type has been established by Chalonge and Divan (1973). It is given in Table 1-10 with other intrinsic parameters for the O stars. The intrinsic gradients have been evaluated according to two hypotheses: (1) that the values of $\Phi_{0\,rb}$ can be determined from the extrapolation shown in Figure 1-11, and (2) that the intrinsic gradient $\Phi_{0\,rb}$ is the same for all O stars and has the mean value 0.73 found by Burnichon (1975) from the study of multiple systems. The final column of Table 1-10 gives the difference, $\Delta\Phi_0 = \Phi_{0\,uv} - \Phi_{0\,rb}$. Its importance is that it can be determined solely from knowledge of the shape of the interstellar extinction law and in a manner practically independent of the hypothesis made concerning the intrinsic values of the gradients (see Divan, 1954, 1970). Its variation is relatively large between types B0 and O5. Therefore, $\Delta\Phi_0$ can serve as an indicator of spectral type and as a test criterion for correct selection of a model atmosphere to represent an O star.

D. Intrinsic Colors of O Subdwarfs

The O subdwarfs are intrinsically faint objects generally observed at quite moderate distances and high galactic latitudes. Many of them are thus practically unaffected by interstellar reddening, and the intrinsic colors are obtained directly from the observed energy distribution. The O subdwarfs are, with some central stars of planetary nebulae, the bluest objects in the sky. For the normal O stars, the bluest $B-V$ actually observed is -0.27 or -0.28 compared to -0.34 for O subdwarfs. Even corrected for interstellar reddening, the $B-V$ colors of the hottest normal O stars (see Table 1-8) are redder than those observed in the case of O subdwarfs.

E. Intrinsic Colors of Wolf-Rayet Stars

Although the dereddening of an observed energy distribution does not imply any hypothesis about its origin, in the background of the usual concept of "intrinsic colors" for Wolf-Rayet stars is the belief that there is a central star with a more or less normal photosphere and that the colors of this photosphere can be observed through the line-emitting region. The continuous opacity of the line-emitting region is discussed in Part III (see pp. 7-74 and 7-75). It can also be observed in V444 Cygni during the eclipse of the O-type primary by the Wolf-Rayet component (Kron and Gordon, 1950; Cherepashchuk and Khaliullin, 1973). The conclusion is that the opacity does not depend on the wavelength and has thus no influence on the colors which are observed. Arguments in favor of a negligible continuous contribution

Table 1-10
Intrinsic Parameters of the O Stars in the BCD System

Sp. Type	D^{a}	s	First Hypothesis $\Phi_0\ rb^{b}$	First Hypothesis $\Phi_0\ uv^{c}$	Second Hypothesis $\Phi_0\ rb^{b}$	Second Hypothesis $\Phi_0\ uv^{c}$	$\Delta\Phi_0$
O4-5	0.035	0.024			0.73	0.48	-0.25
O6-7	0.05	0.038	0.66	0.46	0.73	0.53	-0.20
O8-9	0.06	0.051	0.67	0.51	0.72	0.56	-0.16
O9.5	0.07	0.059	0.68	0.54	0.71	0.57	-0.14
B0	0.085	0.073	0.70	0.58	0.70	0.58	-0.12

[a] D is expressed in dex. The value indicated is that for main-sequence stars. The discontinuities of supergiants are smaller by 0.02 to 0.03 than those of the main-sequence stars of the same spectral type.
[b] The mean wavelength is 5000 Å.
[c] The mean wavelength is 3500 Å.

of the line-emitting region in the visible wavelengths are given by the absence of any observable emission in the Balmer continuum and by the low level of the free-free emission at short wavelengths.

In the case of normal stars, the intrinsic colors are given as a function of the spectral type. This is possible because the classification lines and the continuous spectrum are formed in almost the same layers of the star and have thus a close relationship. For Wolf-Rayet stars, the classification scheme is based on the appearance of the emission-line spectrum which originates outside the photosphere and is probably produced by the deposition of some form of energy not contained in the radiation field of the star. Therefore, no known relation exists between the spectral type of a Wolf-Rayet star and the energy distribution in the continuum of the underlying photosphere. We must be ready to accept the absence of any single-valued relationship.

A rough indication on the colors of the Wolf-Rayet stars is given by the examination of the Balmer region: although strong emission lines may occur near the Balmer limit, especially in the WC spectra, a Balmer absorption larger than about 0.05 or 0.10 dex is excluded in the case of single WR stars. With normal photospheric models, this implies effective temperatures larger than 20000 to 25000 K and hot intrinsic colors.

To have more information than this lower limit to the effective temperature, the main difficulty is not to correct the continuum for interstellar reddening but to determine precisely this continuum. If the continuum is obtained in a domain comprising the Balmer discontinuity, the intrinsic energy distribution of the underlying star can be deduced without much ambiguity for these hot objects from the measurement of the Balmer discontinuity (or from any other method equivalent to this measurement, such as the use of a $(U-B)$ vs. $(B-V)$ diagram and a reddening line) by comparing this discontinuity with that of normal stars of similar absolute magnitude. The intrinsic colors of the underlying star are thus relatively well defined and, when compared with the observed colors in the visible spectral region in which a wavelength-dependent continuous contribution from the line-emitting layers is probably negligible, they give the amount of interstellar reddening. Assuming then a law of interstellar extinction, the whole spectrum can be dereddened and compared with normal stars or with model atmospheres. This comparison either gives an a posteriori justification of the hypotheses which have been made or suggests improvements in the modeling of the photospheric and line-emitting layers and eventually in the dereddening procedure.

All that has been said above applies to single Wolf-Rayet stars. In binaries, when two spectra are seen, the Wolf-Rayet star is almost always associated with an O-type companion which is the brighter star and the main contributor to the observed continuum.

Many attempts have been made to determine the intrinsic colors of Wolf-Rayet stars, either following the lines of thought described above or by an effort to derive the amount of interstellar reddening without any assumption about the continuum observed. From the fluxes which are observable from the Earth, it has not been possible to detect real differences in temperature between the underlying stars of the various Wolf-Rayet subtypes, and the cooler mean temperature found for WC stars by different authors is probably an artifact due to a poorly defined continuum. The work done before 1965 has been summarized by Underhill (1966), and we shall briefly indicate some of the main phases in the research done after this date.

Broadband photometry is clearly not a good tool for analyzing Wolf-Rayet stars, and Pyper (1966), by using spectra with a dispersion of 48 Å/mm, corrected *UBV* magnitudes of Wolf-Rayet stars for the part of the flux due to the emission lines. In the corrected $(U-B)$ vs. $(B-V)$ diagram, the stars have been projected back into the unreddened main-sequence curve with reddening lines taken from Wampler (1961). The systematic difference between WN and WC stars already found in $(U-B)$ vs. $(B-V)$ diagrams uncorrected for line emission (Feinstein, 1964) is smaller in Pyper's corrected diagram, but still exists in the sense that the WC stars are cooler by about 10000 or 20000 K than the WN stars.

Kuhi (1966) determined the energy distribution between 3200 and 11000 Å for quite a few Wolf-Rayet stars with a resolution of 50 Å. Even with this better resolution, line-emission corrections were necessary, and they have been made in order to build a two-color diagram, (λλ 3636–4786), (λλ 4786–5556). In this diagram, the WC and WN stars are grouped around two reddening lines which are very close to each other, with a slight indication of a cooler temperature for the WC stars as found by Pyper (1966), but the effect is much smaller. All stars have been dereddened to $(B-V)_0$ = -0.33, which corresponds to (λλ 4786–5556) = -0.25. Assuming then a standard interstellar reddening law, the whole spectrum has been corrected for interstellar extinction, and intrinsic energy distributions are given for 14 WN and WN + OB stars and 7 WC and WC + OB stars. For the single Wolf-Rayet stars, these energy distributions seemed to have anomalously low color temperatures in the red and infrared part of the spectrum. They have been republished after some corrections by Cohen et al. (1975), together with far-infrared photometry between 1.6 and 11.3 μm. The infrared excesses already seen by Kuhi (1966) are confirmed for Wolf-Rayet stars and attributed to a free-free emission continuum originating in a circumstellar shell and/or to a thermal emission due to circumstellar dust. These emissions are important mostly in the 1- to 10-μm spectral region, but it is not excluded that, at least in some cases, they could extend into the visible part of the spectrum in which the stars have been dereddened; they should be taken into account in the dereddening procedure, but this is a difficult task.

A quite different approach has been adopted by Smith (1968b) who studied Wolf-Rayet stars in the Magellanic Clouds. No a priori hypothesis on the nature of the continuum observed is made, and the stars are dereddened from the knowledge of the amount of interstellar material which is small and can be determined with good accuracy. The visible continuum is defined by two narrowband (~100 Å) filters, b and v, avoiding as far as possible the emission lines (see Table 1-27). Mean $(b-v)_0$ have been calculated for each Wolf-Rayet subtype represented in the LMC. They are given in Table 1-11, together with the individual values.

Table 1-11
Intrinsic Colors of Single WR Stars in the LMC
(from Smith, 1968b)

	$(b-v)_0$	
Spectral Type	Mean Value	Individual Values
WN3	-0.18	-0.12, -0.24
WN4	-0.17	-0.28, -0.14, -0.16, -0.14, -0.13
WN5	-0.14	-0.16, -0.12
WN7	-0.20	-0.16, -0.25, -0.20, -0.20
WN8	-0.15	-0.13, +0.04, -0.35
WC5	-0.21	-0.13, -0.14, -0.18:, -0.29, -0.29

In contrast with the large variation of $(b-v)_0$ among stars of the same subtype, the differences in the mean values from subtype to subtype are small. For all subtypes, the mean $(b-v)_0$ color is that of a late B-type star and cannot correspond to a hot photosphere with a small Balmer discontinuity. This means either that Smith's filters are too wide to avoid reasonably well the emission lines or that the photospheric continuum cannot be seen through the emission layers. The fact that the line strengths can differ by factors up to 10 within each WR subtype (Leep, 1982; Conti et al., 1983b) could explain the large variation in Smith's $(b-v)_0$ colors among stars of the same subtype through a line contribution in the filters, but it could also be an indication that, at least in some cases, no photospheric continuum is seen between the emission lines.

To have a better insight into these problems, it is necessary to measure the energy distributions with the high resolving power which has been obtained only recently thanks to new detectors (Massey, 1982a; Smith and Willis, 1983; Conti et al., 1983b; Vreux et al., 1983; Garmany et al., 1984). The first results on line and reddening-free parameters in the visible continuum (given by Massey, 1982a, 1984) are very puzzling. They cannot be explained by any known photospheric model, and large variations of the interstellar reddening law in the small wavelength range of Massey's observations are very unlikely.

Ultraviolet energy distributions are now available for quite a few Wolf-Rayet stars, and they have confirmed that Wolf-Rayet stars are hot objects with very blue ultraviolet colors and effective temperatures around those of O9 and B0 stars (see, among others, Willis and Wilson, 1978; van der Hucht et al., 1979; Underhill, 1980, 1981; Nussbaumer et al., 1982; Underhill, 1983; Garmany et al., 1984). There was hope that the 2200 Å interstellar feature would give a possibility to deredden the observed energy distributions without any assumption on the intrinsic distributions. It seems now that, in some cases (Nussbaumer et al., 1982) and even in quite a few cases (Garmany et al., 1984), if the extinction corrections which are needed to remove the 2200 Å feature are applied to the observed energy distributions, the resulting $(B-V)_0$ color is bluer than in the case of normal O stars by 0.1 to 0.4 magnitudes. If this is interpreted as an abnormal ultraviolet extinction law, the 2200 Å bump cannot be used to deredden Wolf-Rayet stars. New results on the intrinsic energy distribution of Wolf-Rayet stars should come from precise high-resolution photometry of stars in the Magellanic Clouds, the only region in which the dereddening is not a too serious problem. Nine WR stars have been measured by Smith and Willis (1983) in the Large Magellanic Cloud from 1150 to 7600 Å, with a resolution of 6 to 9 Å between 1150 and 3250 Å, about 2 Å between 3750 and 5020 Å, and about 6 Å in the spectral range 4900 to 7600 Å. The ultraviolet color temperatures range from 29000 to 38000 K and are not correlated with the spectral subtype. They are somewhat bluer than those found by Nussbaumer et al. (1982) for galactic Wolf-Rayet stars dereddened by "nulling" the 2200 Å feature.

F. Spectrophotometric Scans of O and Wolf-Rayet Stars

During the last 20 years, thousands of spectrophotometric scans of stars have been published. The passband is generally 50 Å, somewhat too large to obtain true continua even for O stars, but in certain cases, the passband was reduced to 10 Å for bright stars. These scans are calibrated in terms of absolute fluxes through a wide variety of absolute calibrations. Breger (1976a, 1976b) has made a very careful analysis of this large amount of rather heterogeneous material and has given a catalog of the best scans (937 scans with the corresponding references for 685 stars), all put on the same absolute scale through the calibration of α Lyr by Hayes and Latham (1975). Nine O stars ranging in spectral types from O5 to O9.5 are in the Breger catalog: δ Ori, O9.5 II; ι Ori, O9 III; ζ Ori, O9.5 Ib; S Mon, O7; UW CMa, O7f; ζ Pup, O5f; θ Car, O9.5 V; ζ Oph, O9.5 V; and 10 Lac, O9 V.

Kuan and Kuhi (1976) have published additional scans for 24 O stars using an exit slot of 48 Å in the 3300 to 5100 Å spectral range and 64 Å in the 5100 to 7500 Å range. The fluxes are put on an absolute scale through the use of unspecified secondary standards of Hayes (1970) and of the calibration of Vega by Hayes (1970), which is not significantly different from the calibration by Hayes and Latham (1975) used by Breger. All the O stars of Kuan and Kuhi suffer significant amounts of interstellar reddening. They have been dereddened using the standard reddening law of Whitford (1958) and that A_V which renders the index $E_T = m_{5556} - m_{7530}$ equal to the value which Kuan and Kuhi claim is that of an unreddened star of the same spectral type. The resulting values of A_V are in all cases considerably smaller than the values found when the stars are dereddened using the index $B-V$ instead of E_T. For instance, for the stars 9 Sge, HD 193322, and λ Cep, respectively, Kuan and Kuhi suggest that A_V is 0.75, 0.77, and 1.26 mag, whereas Underhill et al. (1979) find as the result of using three methods that A_V is 0.97, 1.24, and 1.69 mag, respectively. If the scans of Kuan and Kuhi are reliable in the E_T spectral region, their estimates of A_V mean either that the law of interstellar reddening is very abnormal for all their stars or that all the stars observed have intrinsic energy distributions different from those of the normal lightly reddened O stars which are used to obtain the intrinsic values of the color indices E_T and $B-V$. This second alternative has been chosen by Kuan and Kuhi, who conclude that their program O stars have "flat" energy distributions. A systematic error in the longest wavelengths (around 7530 Å) either in the Kuan and Kuhi scans or in the secondary standards used could perhaps explain these results. Another possibility is that all of the stars observed by Kuan and Kuhi have a detectable amount of excess radiation near 7530 Å. In any case, the reality of "flat" energy distributions should be tested also with unreddened or nearly unreddened stars before concluding that the energy distributions of the O and Of stars are truly indicative of atmospheric extension.

Spectrophotometric scans between 3300 and 11100 Å have been published by Cohen et al. (1975) for 32 Wolf-Rayet stars, with an exit slit of 50, 96, or 128 Å. Twenty of these stars are measured in the whole spectral range and 12 from 8000 to 11100 Å. Infrared photometry between 1.6 and 11.3 μm is also given for 23 stars. Among them, 17 have scans which can be combined with the infrared photometry to produce composite energy distributions. The observed and the dereddened fluxes are tabulated. This material has been used by the authors to discuss the infrared excesses of the Wolf-Rayet stars.

Vreux et al. (1983) have given energy distributions between 5750 and 10350 Å for 26 Southern Hemisphere Wolf-Rayet stars in the form of tracings with a resolution of about 8 Å.

G. Ultraviolet Intrinsic Energy Distributions

The absolute ultraviolet energy received at the Earth has been measured for a number of O stars by means of spacecraft spectrophotometric and photometric instruments which have been listed above, in the section *Main Ultraviolet Programs*. A detailed comparison of OAO-2, ANS, and TD1-S2/68 observations by Gilra et al. is still in preparation, but Koornneef et al. (1982) gave a structured presentation of the data contained in these various source catalogs. Magnitudes versus wavelength for 531 individual stars (including 6 WR and 39 O stars) are represented in a homogenized absolute calibration system. The OAO-2 photometry is given for the whole sample; it is supplemented by OAO-2 spectrophotometry of 213 stars, ANS photometry for 363 stars, and TD1 fluxes for 488 objects. For early-type stars, the agreement between the various ultraviolet photometric systems is generally better than 0.10 mag.

In order to get intrinsic energy distributions, the observed data must be corrected for the wavelength-dependent extinction due to the interstellar medium which affects all O stars. The shape of the interstellar extinction curve through the ultraviolet has been deduced by Nandy et al. (1975, 1976) from TD1-S2/68 spectrophotometric scans of several hundred O- and B-type stars distributed throughout the galactic plane. Savage and Mathis (1979) have rediscussed all the information available on the average interstellar extinction curve and have provided a useful table of $A_\lambda/E(B-V)$ extending from $\lambda^{-1} = 0$ to $10\ \mu m^{-1}$ (see their Table 2). The existence of differences of interstellar extinction in different areas of the Galaxy is still a controversial question. It must be noted that, in the present state of our knowledge, our nearest neighbors, the Magellanic Clouds, have mean interstellar extinction curves similar to the galactic one from the infrared to about 2500 Å but sensibly different from it at shorter wavelengths: smaller hump at 2200 Å and higher far-ultraviolet extinction for the LMC (Nandy et al., 1981; Koornneef and Code, 1981), no hump at 2200 Å and still higher far-ultraviolet extinction for the SMC (Prévot et al., 1984).

We shall now review the major sources of information for the intrinsic ultraviolet energies of O stars, bearing in mind that: (1) an interstellar extinction correction has been applied from the evaluation of $E(B-V)$ and the use of an average galactic extinction law, and (2) the absolute calibrations of different experiments are somewhat connected and are often based on the flux for η UMa as defined in Bohlin et al. (1980), which these authors consider accurate to within 10 percent longward of 1200 Å.

From TD1-S2/68 observations, Nandy et al. (1976) provided a grid of intrinsic colors $(U_\lambda - V)_0$ from 1400 to 2740 Å with a mean rms error of ± 0.1 mag for O and B stars. More recently, Carnochan (1982) presented a study of intrinsic colors of stars O to G, utilizing the full S2/68 data set and extending the work of Nandy et al. Different ways for calculating the $(U_\lambda - V)_0$ were used, and the arithmetic mean was adopted as the intrinsic color (mean error varying between 0.05 and 0.4 mag). These intrinsic colors have been corrected from the absolute calibration of S2/68 (whose error is estimated at less than 20 percent in the absolute fluxes) to the standard calibration of Bohlin et al. (1980) for the star η UMa. The results are given in Figure 1-12 by a dotted line. The arrows indicate the Δm corrections introduced by using the standard calibration of Bohlin et al. (1980). One can see in Figure 1-12 that for O stars the two sets of TD1-S2/68 intrinsic colors differ appreciably, probably because of different dereddening, normalizing, and smoothing procedures. Llorente de Andrés et al. (1981) also made a study of ultraviolet intrinsic colors of early-type stars from TD1-S2/68 data and compared them with the

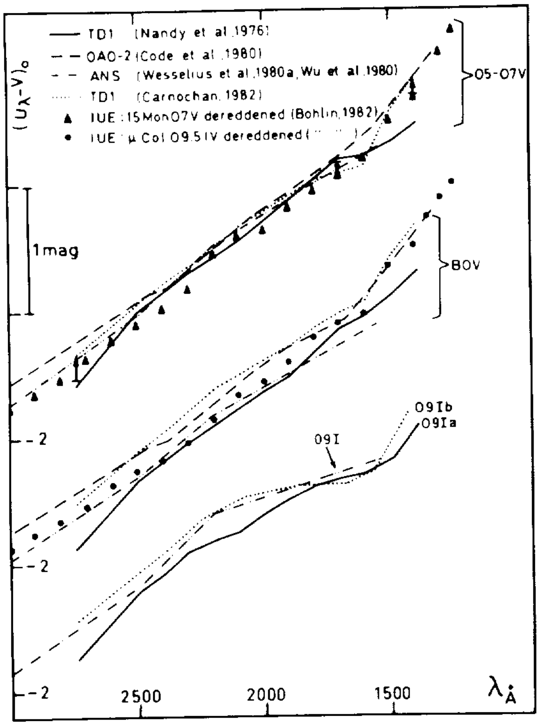


Figure 1-12. Comparison of different determinations of ultraviolet intrinsic colors for O stars.

predictions of stellar model atmospheres adopting the effective temperature scale described by Castelli et al. (1980). Their results are in agreement with those of Nandy et al. (1976), and they conclude also that the dwarfs and giants of the same spectral type have the same relative flux distribution, while the supergiants have redder $(U_\lambda - V)_0$ colors. It can be seen that the difference between the flux distributions published for dwarfs and supergiants cannot be removed by changing the values of $(B-V)_0$ used to deredden the dwarfs and the supergiants. Moreover, comparison with the ultraviolet energy distributions predicted by stellar model atmospheres shows that dwarf and giant stars of spectral type O9 to B2 are redder than the models. This can be interpreted by the underestimation for the models both of line opacities and of the adopted microturbulence velocity. For supergiants and especially the hottest ones, a similar conclusion is reached.

Intrinsic ultraviolet colors based on ANS data (from 1550 to 3300 Å) have been presented by Wesselius et al. (1980a) from the study of 182 main-sequence stars and by Wu et al. (1980) for stars of all luminosity classes. Wesselius et al. considered stars with small reddening, but O stars are included up to $E(B-V) = 0.30$. They assume a gradual change of each color index with effective temperature and smooth the original determinations with a polynomial fit. They find that, for the hot stars, the spread in each of the ultraviolet colors is smaller than 0.02 mag. Wu et al. do not smooth the original determinations; their results are in reasonable agreement with those of Wesselius et al. for hot main-sequence stars. A careful comparison of the observed color indices $(U_\lambda - V)_0$ with Kurucz (1979) models has been made by Wesselius et al. (1980a). They discuss the data on main-sequence stars and also on 56 giants and supergiants using, as representation of the data, color index versus effective temperature diagrams. They confirm that supergiants earlier than B6 are much redder than main-sequence stars at the same spectral type. Admitting that the absolute calibrations of the ANS magnitudes have independent errors, Wesselius et al. adjust the ANS-observed colors in order to fit the model colors for main-sequence B stars. In these conditions, the observed O stars are slightly bluer than the models. Galecki et al. (1983) redetermined intrinsic ultraviolet colors of early-type stars from ANS data, assuming that there is a linear relation between the ultraviolet color excesses and $E(B-V)$; a comparison to the model values of Wesselius et al. has been made, but without any adjustment; it confirms that, in the ANS photometry, the hottest dwarfs are slightly bluer than the models, but only for wavelengths longer than 1800 Å.

Code et al. (1980) presented a catalog of magnitudes from OAO-2 filter photometry (1330 to 4250 Å) for 531 stars and tabulated also mean values of $(U_\lambda - V)_0$ for nine main-sequence types among which are O7 and B0.

No published tables of intrinsic colors from IUE data (1200 to 3100 Å) are available up to now. However, for comparison, we plot in

Figure 1-12 the individual values of colors for two dereddened O stars from Bohlin (1982).

The results from different authors agree to within ± 0.1 mag for $\lambda > 1600$ Å (see Figure 1-12). At shorter wavelengths, the agreement may be as good if the Carnochan (1982) values deduced from the TD1 data are preferred to those of Nandy et al. (1976); this better agreement with OAO-2 and IUE results is not due only to the recalibration of TD1-S2/68 on η UMa absolute fluxes of Bohlin et al. (1980).

Finally, one can say very briefly that, without considering any effective temperature scale for stars, when Kurucz (1979) models grossly match the ultraviolet observations above about 2000 Å, they become progressively bluer than the observations at shorter wavelengths.

IV. EFFECTIVE TEMPERATURES AND BOLOMETRIC CORRECTIONS

In this section, we shall review the main determinations of effective temperatures of O and Wolf-Rayet stars. Bolometric corrections are reviewed at the end of the section.

A. Normal O Stars

Effective temperatures of O stars have been determined mainly by three independent methods: from the line spectrum by comparison to the predictions of model atmospheres, from the Zanstra principle applied to the surrounding H II region of an exciting star, and from the integrated flux.

The first method consists of comparing the observed line equivalent widths or profiles to the predictions of model atmospheres. This was done by Conti (1973) using his measurements of O star line strengths and the non-LTE model computations of Auer and Mihalas (1972). Comparison of the predicted and observed runs of the line ratio, λ 4471 He I/λ 4541 He II, used for spectral classification led to the scale of effective temperatures which is shown in Figure 1-13 by solid lines. The estimated probable error due to measuring uncertainties is ± 500 K at a given spectral type. Uncertainties originating in the linking of real stars to the models are to be added to this.*

Kudritzki and his coworkers (Kudritzki, 1980, 1981; Kudritzki et al. 1982; Simon et al., 1983) performed a detailed non-LTE analysis of five individual earliest O stars from high dispersion (12 or 2.6 Å/mm) spectra with low noise and high contrast. Equivalent widths of the He I λ 4471 line, together with the observed profiles of hydrogen and ionized helium lines, are used for the determination of effective temperature, gravity, and helium abundance. The results for the five program stars are compatible with a normal helium abundance, and they lead to the values of effective temperatures and gravities which are given in Table 1-12.

Until the last decade, fluxes over a large wavelength range were not directly observed, and Hjellming (1968) and Morton (1969) estimated the temperatures of O-type stars which excite H II regions by means of the Zanstra method. The Zanstra principle states that all Lyman continuum photons are ultimately converted to Balmer line or continuum photons; measurement of the radio free-free emission and/or the Hα emission in the surrounding H II region of an O star permits one to estimate the Lyman continuum flux from the star. One compares this to the predictions from model stellar atmospheres and deduces T_{eff}. Two problems arise with this method: if the nebula is density-bounded rather than ionization-bounded or if it contains dust, the temperature of the star is underestimated; if another undetected star excites the nebula, the temperature is then overestimated. Both Hjellming and Morton took great care to use nebulae which seem free from such effects. Morton's effective temperature scale for O stars is shown in Figure 1-13. The accuracy is estimated to be ± 10 percent.

Recently, fluxes from λ 1100 Å to the infrared have become accessible to observation

*See also Abbott and Hummer (1985) for the effects of "wind blanketing" on T_{eff} deduced from a line ratio.

Table 1-12
Effective Temperature of O Stars

HD or BD	Sp Conti	Sp Walborn	Conti Scale	Kudritzki T_{eff}	Kudritzki log g	Morton Scale	Code et al.	UDPD[a] Underhill	Pottasch et al.
93129	O3f	O3If*	(52000)	45000 ± 2000	3.60 ± 0.15			46000	
93128	O3((f))	O3V((f))	(55000)	48000 ± 3000	3.85 ± 0.15			(63000)	
303308	O3	O3V((f))	(55000)	45500 ± 3000	3.90 ± 0.15			50100	
93205	O3	O3V	(55000)					(59000)	
93250	O3	O3V((f))	(55000)	52500 ± 2500	3.95 ± 0.15			50000	
15570	O4f	O4If+	(47500)					44000	
16691	O4f	O4If+	(47500)					(58000)	
66811	O4ef	O4I(n)f	(47500)	41000 ± 2000	3.50 ± 0.15		32510 ± 1930	35000	
46223	O4((f))	O4V((f))	(50000)					34800	43600
164794	O4((f))	O4V((f))	(50000)					54000	44700
168076	O4((f))	O4V((f))	(50000)					56000	
96715	O4V[b]	O4V((f))	(50000)					31900	
242908	O4[b]		(50000)					50100	
15629	O4((f))	O5V((f))	(50000)					46200	
14947	O5f	O5If+	(44500)			48000		28000	
+54°2761	O5f[b]		(44500)			48000		44000	
15558	O5(f)	O5III(f)	(46000)			48000		46200	
93843	O5III[b]	O5III(f)var	(46000)			48000		(61000)	
93204	O5((f))	O5V((f))	(47000)			48000		(61000)	
97484	O5[b]		(47000)			48000		24800	
193682	O5[b]		(47000)			48000		34900	
93403	O5.5(f)	O5III(f)var	(43500)					38000	
192281	O5.5(ef)	O5Vn((f))p	(43500)					29800	
168112	O5.5((f))	O5III(f)	(44500)					59000	
46150	O5.5((f))	O5V((f))	(44500)					35000	43600
210839	O6ef	O6I(n)fp	(40000)			40000		37500	
206267	O6	O6.5V((f))	42000			40000			36200
+60°2522	O6.5IIIef	O6.5(n)(f)p	39000						33500
152723		O6.5III(f)	39000						36800
42088	O6.5V	O6.5V	40000						36800
48099	O6.5V	O7V	40000					37000	
24912	O7.5I	O7.5III(n)((f))	35500						32500
188001	O8If	O7.5Iaf	34500			33500		34500	
151804	O8If	O8Iaf	34500			33500		34500	
47839	O8III((f))	O7V((f))	35500			33500		35500	31400
36861	O8III((f))	O8III((f))	35500			33500		35000	31300
214680	O8III	O9V	35500			33500		35500	
57060	O8.5If	O7Ia:fpvar	33500					33500	
193322	O8.5III	O9V:((n))	34500					34000	
57061	O9I	O9II	33000			32000		33000	
149757	O9V(e)		34500			32000	31910 ± 2040	34000	
36512	O9V[b]	B0V	34500			32000		34500	
30614	O9.5I	O9.5Ia	31500			31000		25000	
188209	O9.5I	O9.5Iab	31500			31000		31000	
36486	O9.5I	O9.5II	31500			31000		31000	
37742	O9.5I	O9.7Ib	31500			31000	29910 ± 2110	27500	
34078	O9.5V	O9.5V	33000			31000			29500
37468	O9.5V		33000			31000		31500	

[a]Underhill et al. (1979).
[b]Spectral type from Humphreys (1978).

for a number of stars, and it is now possible to determine effective temperatures based mainly on observational data, although still depending in part on theoretical model atmospheres.

By definition, the effective temperature T_{eff} of a star is given by:

$$\sigma T_{eff}^4 = \int_0^\infty \mathcal{F}_\lambda(*)\, d\lambda \quad , \qquad (1\text{-}9)$$

where σ is the Stefan-Boltzmann constant, and $\mathcal{F}_\lambda(*)$ is the monochromatic emergent flux at the stellar surface in ergs s^{-1} Å^{-1} cm^{-2}. The monochromatic emergent flux is related to the monochromatic flux measured outside the Earth's atmosphere and corrected for interstellar extinction, f_λ, by means of the angular diameter, θ_λ:

$$\mathcal{F}_\lambda(*) = 4 f_\lambda / \theta_\lambda^2 \quad . \qquad (1\text{-}10)$$

Assuming that θ_λ has the same value at all wavelengths, it is possible when this value is known to find the effective temperature, estimating only from theory the nonobserved part of the integrated flux:

$$\sigma T_{eff}^4 = 4\,\theta^{-2}\, G(\text{model}) \int_{\lambda_1}^{\lambda_2} f_\lambda\, d\lambda \quad ,(1\text{-}11)$$

with $G(\text{model}) = 1 +$

$$\left(\int_0^{\lambda_1} F_\lambda\, d\lambda + \int_{\lambda_2}^{\infty} F_\lambda\, d\lambda \right) \Big/ \int_{\lambda_1}^{\lambda_2} F_\lambda\, d\lambda \quad . \qquad (1\text{-}12)$$

In Equation (1-12), λ_1 and λ_2 are the lower and upper limits of the observed wavelength range and F_λ is the monochromatic flux of the model.

This method has been used primarily by Code et al. (1976) for three bright O stars and by Underhill et al. (1979) and Underhill (1982) for 18 mostly late O and 24 early O stars, respectively. The fluxes were observed from the ground and from OAO-2, TD1-S2/68, or IUE

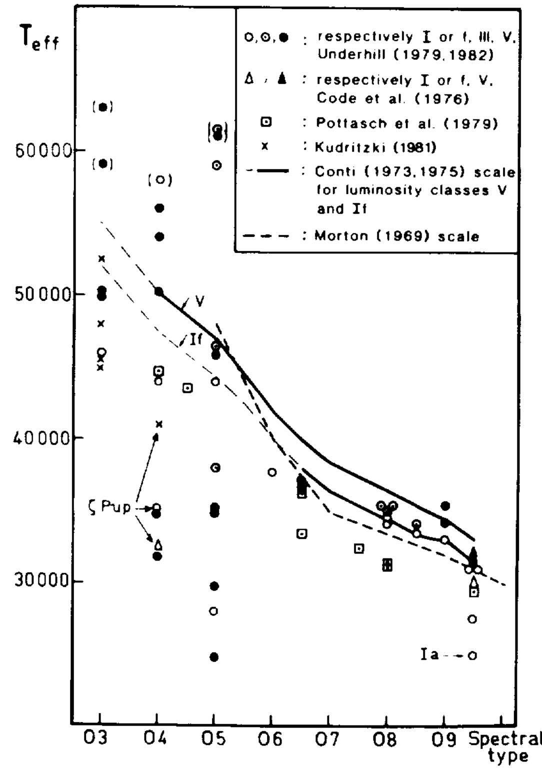


Figure 1-13. The effective temperature versus spectral type for O stars.

in the ultraviolet range, and the model-dependent nonobserved part of the flux is not too large.

Code et al. used the angular diameters measured at λ 4430 Å by Hanbury Brown et al. (1974) by means of the intensity interferometer, and their effective temperatures can be considered as purely empirical. Unfortunately, the results by Code et al. (1976) are very few in number for O stars. However, ζ Pup was considered, and it is given a very low temperature compared to the previous scale by Morton or by Conti (see Figure 1-13); this surprising result has been discussed at length in several papers. The uncertainty in T_{eff} is estimated to be about 7 percent for the O stars of Code et al.

Underhill et al. (1979) and Underhill (1982) used angular diameters found by replacing $\mathcal{F}_\lambda$ in Equation (1-10) by the monochromatic fluxes from a model atmosphere in the infrared or the visible range, following the suggestion of

Blackwell and Shallis (1977). The late O stars show a trend which is compatible with the previous scales, except for the O9.5 supergiants which have systematically lower temperatures (Figure 1-13). For the early O stars, the effective temperatures range from 25000 to 63000 K (see Figure 1-13). Since the uncertainty in these results was estimated to be about 5 percent, this range was considered as probably real even though the highest effective temperatures may be suspect. This would indicate that a spectral type of O5, O4, or O3 cannot be correlated closely with effective temperature. The median effective temperature at each early O spectral type is as follows: $\langle T_{eff} \rangle = 42300 \pm 12300$ K at O5; $\langle T_{eff} \rangle = 45500 \pm 9800$ K at O4; and $\langle T_{eff} \rangle = 53600 \pm 6300$ K at O3. It increases as the spectral type becomes earlier, but the standard deviation of each result is large enough for types O5 and O4, greatly exceeding the expected typical error of a single result, to suggest that all of the stars which are given spectral type O5 or O4 do not have about the same effective temperature. At type O3, the spread about the mean value is not very much larger than what can be expected from the errors inherent in the method. However, because the sample is small at this subtype, consisting of only five stars, it is uncertain whether or not the stars given type O3 truly have about the same effective temperature. These median values are below the Conti (1973, 1975) scale for O4 and O5 types and in agreement with it for the O3 type, while the four values found by Kudritzki and his coworkers (see above) for O3 stars are below the Conti scale. In addition, the early O-type stars having the "f" characteristic lie mainly in the lower part of Figure 1-13, but one cannot say that they are systematically cooler than the early O-type stars which do not have this characteristic.

The results by Pottasch et al. (1979), based on a study of ten O-type exciting stars of diffuse nebulae which considers all the observed flux shortward of the Lyman limit, lie below the Conti scale, parallel to it, and are roughly in qualitative agreement with those of Underhill (see Figure 1-13).

Individual results on T_{eff} determinations are given in Table 1-12, along with the T_{eff} scales of Conti and Morton.

One can say in summary that, for the early O stars, the effective temperatures derived by the method of the integrated flux do not agree well in general with the effective temperatures deduced from the analysis of the line spectrum. The observed true dispersion in the flux effective temperatures is very large. This would mean that the strengths of the classification lines used for spectral-type determinations are not closely related to the effective temperatures in a unique monotonic manner. For the sample of early O stars considered by Underhill (1982), there is no apparent correlation of the flux effective temperature with spectral type or with luminosity. In the case of late O stars, the agreement is relatively good between the two methods, but for supergiants of type O9.5, the flux effective temperatures are significantly lower than the effective temperatures of Conti based on the relative strengths of He I and He II lines.

Balmer discontinuities, D, have been measured by L. Divan (unpublished data) in the BCD system for quite a few O stars. In the case of late O stars (O7 to O9), the values of D are well correlated with the spectral subtype and the luminosity class which are deduced from the analysis of the line spectrum (see Figure 1-2). For the early O stars (O3 to O5), the Balmer discontinuities are very small and cannot exclude a rather large range in high effective temperatures; however, the observed D values are always smaller than for the O7 to O9 subtypes, and if the Balmer continuum is purely photospheric, effective temperatures lower than 35000 K seem to be excluded.

The discrepancy between the flux effective temperatures and the line effective temperatures in the case of O stars raises the question of the validity of using empirically selected classification lines to determine the conditions in the photosphere. These lines are probably partly formed in the outermost regions of the atmosphere (Underhill, 1982) in which conditions are probably determined more by the deposition of nonradiative energy and momentum than by the heating provided by the passage of radiation from the center of the star. This important point will be discussed in Part III.

B. O Subdwarfs

At the time of writing, no direct determination of the effective temperature from observed fluxes and angular diameters has been made for hot subdwarfs. Fluxes have been measured in a large wavelength range, from the infrared to about 1200 Å in the ultraviolet part of the spectrum, and the models available are probably good enough for estimating the unobservable amount of energy. The difficulty is that the hot subdwarfs with their low luminosities and high surface brightness can have only very small radii, and even at quite moderate distances, their angular diameters are unmeasurable with the present observing techniques. In fact, even if the angular diameters were measurable, the so-called "direct" determination of the effective temperature in the case of these very hot stars would be somewhat "model-dependent" through the unobservable part of the flux which sometimes exceeds the observable part. The benefit of using measured integrated fluxes from 1200 Å to some long wavelength is then less than in the case of cooler stars. Another difficulty is that the chemical composition of the O subdwarfs is a priori unknown, and the possibility of using angular diameters deduced from model atmospheres, following the method applied to the O stars of Population I by Underhill et al. (1979) and Underhill (1982), has not yet been explored. For the O subdwarfs, the effective temperatures are derived by a more indirect method.

The indirect method consists of: (1) observing part of the energy distribution of the star in the continuum and/or in the lines, and (2) calculating a model atmosphere for which the energy distribution matches the observed features. If the match is good, one accepts that the effective temperature of the star is equal to the effective temperature of the model. This is true only if the physics of the model truly represents what happens in the atmosphere of the star. Otherwise, even a good match could be completely misleading. The probability that the physics is good increases if many features, belonging to different elements with different degrees of ionization and spread over a large spectral range, are correctly represented. Thus, the different attempts to determine the effective temperature differ by the observed features used (nature, number, spread in wavelength, lines or continuum, etc.) and by the physics of the model.

In the case of massive O stars, the chemical composition is largely solar, and the grid of useful models depends only on two parameters: the effective temperature and the gravity.* The situation is much more complicated for hot subdwarfs because they have a large variety of chemical compositions, from pure helium to almost pure hydrogen (see Table 1-13), and the grid of models must include parameters of chemical composition.

In the atmospheres of the hot stars, non-LTE effects are important not only for line profiles but also for the continuum. For a given temperature, the non-LTE effects decrease as gravity increases, and the hot subdwarfs are less influenced than the main-sequence stars. However, important non-LTE effects exist in many hot subdwarfs and Kudritzki (1976) showed that the difference between LTE and non-LTE disappears only for $\log g = 7$ at $T_{eff} = 45000$ K and $\log g = 6$ at $T_{eff} = 40000$ K. These limits depend on the chemical composition and are given here for $N(\mathrm{He})/N(\mathrm{H}) = 0.1$.

From the time of their identification as a spectroscopic group with very blue colors, broad hydrogen lines, and the presence of ionized helium, the O subdwarfs were known to have a high effective temperature. However, the first quantitative determinations were performed only much later by Searle and Rodgers (1966), Sargent and Searle (1968), Newell et al. (1969), Greenstein (1971), and Greenstein and

*If the wind backscattering discussed by Abbott and Hummer (1985) is not taken into account.

Table 1-13
Atmospheric Parameters of O Subdwarfs

Star	$T_{eff}(10^3$ K)	Log g	$\frac{N(\mathrm{He})}{N(\mathrm{He}) + N(\mathrm{H})}$	References[a]
HD 49798	47.5	4.25	0.5	1
BD + 75° 325	55.0	5.3	0.6	2
SB 21	36.3	5.4	1.0	3
SB 58	38.0	4.5	0.7	3
LB 1566	42.7	5.0	0.2	3
SB 705	44.7	5.8	0.5	3
SB 933	49.0	5.5	0.4	3
LS 630	55.0	6.0	0.5	3
SB 884	39.8	6.1	0.12	3
SB 169	37.2	6.3	0.04	3
EG 55	38.0	5.8	0.02	3
SB 38	38.9	5.4	0.03	3
TON S-103	39.8	6.5	1.0	3
HD 127493	42.5	5.25	0.6	4
HD 149382[b]	35.0	5.5	0.04	5

[a] References: 1. Kudritzki and Simon, 1978; 2. Kudritzki et al., 1980; 3. Hunger et al., 1981; 4. Simon, 1982; 5. Baschek et al., 1982.
[b] The spectral type of this star is sdOB.

Sargent (1974). The effective temperature was deduced from the reddening free parameter $Q = (U-B) - 0.72\ (B-V)$ through a relation between Q and T_{eff} established with model atmospheres having a solar chemical composition.

Baschek and Norris (1975) have analyzed the O subdwarf, HD 149382, by improved methods, using monochromatic measurements of the continuum instead of broadband colors, detailed profiles of H, He I, He II lines, and a grid of models having gravities as high as log $g = 6.5$. The LTE assumption is made, but non-LTE effects are discussed. It was clear that more realistic models were necessary to interpret the physical conditions in the atmospheres of hot subluminous stars.

Such models were calculated at Kiel by Kudritzki, who began to publish grids of non-LTE models with high gravities, high temperatures, and different chemical compositions (Kudritzki, 1976). He investigated the effects of non-LTE situations in hot subdwarfs and gave the first applications of the new models to observations. Almost at the same time, the first observations by Wesselius (1977) of O subdwarfs with the ANS Satellite extended the observed domain to the far ultraviolet. Wesselius compared the ultraviolet energy distributions corrected for interstellar absorption using the 2200 Å feature with Kudritzki's models for three O subdwarfs, BD + 28° 4211, BD + 75° 325, and HD 127493. The energy distributions were explained by high temperatures and high gravities, but a more complete set of models and of observations are necessary to draw any further conclusion. At these very high temperatures, continuum energy distributions only are not sufficient.

The first very complete analysis of an O subdwarf was published by Kudritzki and Simon (1978) for HD 49798. Well-observed equivalent widths and profiles for a large number of lines of H, He I, and He II were analyzed by means of a grid of non-LTE models sufficiently dense in effective temperature, gravity, and helium abundance. The fit is made with the equivalent widths, and although the number of analyzed lines greatly exceeds the number of parameters of the model, a solution in T_{eff}, log g, and N(He)/N(H) can be found. With this solution, complete line profiles, together with the energy distribution in the continuum, were calculated and compared to the observations. The fit is excellent in all cases, even for the He I line profiles which are completely unexplained by LTE models. Most probably, the physics of Kudritzki's models is very close to what really happens in the atmospheres of the hot subdwarfs.

These models are now used by the Kiel group to analyze more O subdwarfs: BD + 75° 325 (Kudritzki et al., 1980); HD 149382 (Baschek et al., 1980); SB 21, SB 58, LB 1566, SB 705, LS 630, SB 933, SB 884, SB 169, EG 55, SB 38, and TON S-103 (Hunger et al., 1981); HD 127493 (Simon, 1982); and HD 149382 (Baschek et al., 1982). In some of these analyses, the far-ultraviolet He II line at 1640 Å is measured on high dispersion IUE spectra and introduced into the fit diagrams.

Other groups have started to analyze hot subdwarfs, generally using LTE models for the continuum and sometimes non-LTE models for the lines (see, for instance, Wesemael et al., 1982). See also Caloi et al. (1982) for the very interesting case of an O subdwarf (presence of the He II lines 4686, 4542, and 1640 Å) which was discovered by Searle and Rodgers (1966) in the globular cluster NGC 6397. Membership in the cluster seems well established, and the star is situated above and to the left of the horizontal branch. Unfortunately, the effective temperature and the gravity of this star are still not well known.

The results of some of the best analyses of O subdwarfs are given in Table 1-13. The effective temperatures are very high, and although O subdwarfs cannot be divided into spectroscopically defined subtypes, their range in temperature is almost as large as the entire interval covered by the normal O stars. The hottest O subdwarfs have effective temperatures of about 55000 K, like those of O3 Population I stars. However, the extremely high temperatures of some central stars of planetary nebulae (Table 1-14) are not observed for the O subdwarfs. (We must note that the temperatures of Table 1-14 were calculated by Mendez et al. (1981) using exactly the same method as for the O subdwarfs of Table 1-13.)

C. Wolf-Rayet Stars

The classification scheme of Wolf-Rayet stars, based on the visible emission-line spectrum, describes an ionization/excitation temperature which ranges from $\simeq 50000$ to $\simeq 100000$ K or so. The problem is how to relate (if there is a relation) the high electron temperatures in the emission-line forming region to the thermal effective temperature which describes the total flux of radiation flowing through the atmosphere.

Morton (1970, 1973) obtained effective temperatures for Wolf-Rayet stars by making a Zanstra analysis (see preceding section) of the observed radio emission from ring nebulae surrounding some WN stars. His results for individual stars are listed in Table 1-15.

Observations covering a wide wavelength range have become available with the advent of sensitive ultraviolet experiments, and recent determinations of Wolf-Rayet temperatures based on ultraviolet data have been carried out. Willis and Wilson (1978) observed the ultraviolet spectra of nine WR stars with S2/68 on TD1 and compared the flux distributions over the range 1350 Å to 1 μm, corrected for interstellar extinction, with blackbodies and model atmosphere calculations to yield temperatures. These results are listed in Table 1-15, where the mean values of T_k (comparison with Kurucz et al. (1974) models) and T_b (comparison with blackbody distributions) are given. The results lie in the neighborhood of 30000 K; it should be noted that the Zanstra temperatures derived in the same study by Willis and Wilson from the He II λ 1640 emission line are quite similar to the temperatures T_k and T_b.

Table 1-14
Atmospheric Parameters of Central Stars of Planetary Nebulae
(from Méndez et al., 1981)

Planetary Nebula	T_{eff}(10^3 K)	Log g	$\frac{N(\mathrm{He})}{N(\mathrm{He}) + N(\mathrm{H})}$
NGC 1360	60	5.25	0.03
1535	50	4.5	0.09
4361	70	5.0	0.09
7293	65	6.5	0.01
Abell 7	75	7.0	0.01
Abell 36	65	5.2	0.13

Underhill (1980, 1981, 1983) used the integrated flux method as explained in the preceding section on O stars to estimate the effective temperatures of ten Wolf-Rayet stars. The data came from TD1-S2/68, ANS, IUE, and ground-based photometry. In general, the results agree well with those of Willis and Wilson, as can be seen in Table 1-15.

Nussbaumer et al. (1982) analyzed low-resolution IUE spectra of 15 Wolf-Rayet stars. They deduce color excesses for each star from the observed strength of the 2200 Å feature and compare the combined dereddened ultraviolet/visible continua with blackbody energy distributions to derive color temperatures. They also deduce temperatures from a Zanstra analysis of the He II emission-line spectra.

If we except these last values, as well as some isolated results by Morton (1970, 1973), the agreement between the different determinations is fairly good; it seems that the effective temperatures of Wolf-Rayet stars lie in the range 25000 to 35000 K, mainly around 30000 K; there appears to be no definite difference in temperature between sequences and subclasses.*

For the WN and WC stars of the higher excitation subclasses, Nussbaumer et al. (1982), in view of the existence of pronounced extended atmosphere effects in the continuum, do not infer the value of T_{eff} from the continuum data, and they adopt the Zanstra temperatures. The last column of Table 1-15 gives their adopted values of T_{eff}, with the Zanstra temperatures in parentheses. In three cases, the Zanstra determination was not reliable, and a value of T_{eff} similar to that found for stars of the same excitation has been adopted (double parentheses). If we consider these results, there is a trend to higher effective temperatures ($\approx$ 40000 K) for the WC5 and the early WN subtypes.

*cf. Chapter 3 for a critique of all these results.

D. Bolometric Corrections

For normal O stars, the bolometric corrections, BC, can be estimated from the tabulation of Buser and Kurucz (1978), which gives the calculated values of bolometric corrections, BC_{calc}, as a function of T_{eff}, log g, and abundance for the different models of Kurucz (1979).

The appropriate value of bolometric corrections are the "BC'" values of Buser and Kurucz:

$$BC' = BC_{calc} + 0.100 ,$$

which represent the best fit of the theoretical calculations to the Code et al. (1976) empirical results.

Table 1-15
Temperature Estimates for WR Stars

HD	Sp. (van der Hucht et al., 1981)	Morton (1970, 1973)	Willis & Wilson (1978)	Underhill (1983)	Nussbaumer, et al. (1982)
9974	WN3 + abs				39000
56925	WN4	53600			
147419	WN4	42000			
187282	WN4				((41000))
219460	WN4.5	48000			
50896	WN5	29500	29700	29900	(41000)
89358	WN5	54200			
191765	WN6		32100	24800	(41000)
192163	WN6	30700	31500	39900	34000
211853	WN6 + O	28600			
151932	WN7			25000	24000
92740	WN7 + abs			25100	30000
93131	WN7 + abs			29700	30000
86161	WN8				30000
96548	WN8	25000		24900	32000
16523	WC5				((41000))
165763	WC5		31000	30000	((41000))
113904	WC6 + O9.5I		31000		
156385	WC7		28700		30000
193793	WC7 + abs		28000		
192103	WC8		29000	25200	26000
68273	WC8 + O9I		30500		
164270	WC9				22000
184738*	WC9			18000	

*Central star of a planetary nebula.

Table 1-16 gives bolometric corrections *BC* (= *BC′* values) valid for stars having normal solar abundance and T_{eff} and log *g* value pertinent to O stars.

In the case of O subdwarfs, the bolometric correction has been estimated by Rossi (1979) using grids of model atmospheres in wide gravity and chemical composition ranges (log *g* up to 5.5, *X* from 0.9 to 0.1, and *Z* from $Z_\odot$ to $Z = 10^{-5}$). The result is that, within a few hundredths of magnitude, the ($BC/\log T_{eff}$) relation does not depend on gravity or on chemical composition. The bolometric corrections of Rossi as deduced from his Figure 5 at log $g = 5$ are very close to those of Buser and Kurucz (1978) for the same gravity, and the values of Table 1-16 at log $g = 5$ are a fairly good approximation for O subdwarfs.

V. ABSOLUTE VISUAL MAGNITUDES

In this section, we shall review the principal determinations of the absolute magnitudes of O and Wolf-Rayet stars.

A. Normal O Stars*

The general problem of determining the absolute magnitudes of O-type stars has been discussed in detail by Underhill (1966). A few

*cf. corresponding discussion in Chapter 3.

Table 1-16
Bolometric Corrections for O Stars

T_{eff} log g	BC (mag)	T_{eff} log g	BC (mag)	T_{eff} log g	BC (mag)
25000		35000		45000	
3.5	-2.47	3.5	-3.28	4.5	-4.03
4.0	-2.50	4.0	-3.26	5.0	-4.02
4.5	-2.53	4.5	-3.28		
5.0	-2.55	5.0	-3.30		
30000		40000		50000	
3.5	-2.86	4.0	-3.66	4.5	-4.37
4.0	-2.91	4.5	-3.65	5.0	-4.37
4.5	-2.95	5.0	-3.65		
5.0	-2.97				

points are recalled here briefly. The O stars are few in number, and they are distributed in the galactic plane; not one of them is sufficiently close to have a trigonometric parallax. In addition, since they are reddened by interstellar material, it is necessary to correct their radiation for the wavelength-dependent extinction due to interstellar dust.

Most of the determinations of absolute magnitude have been obtained using O stars belonging to stellar groups (multiple systems, clusters, and associations) for which the distance and the average extinction are known from other stars, often main-sequence B stars, independently of any hypothesis about the structure of the Galaxy. This method is used in order to protect for O stars their important role as indicators of distance in the Galaxy. It is necessary, of course, to establish from the start that the star under study is indeed a physical member of the group, multiple system, or cluster being considered.

Therefore, the method of cluster fitting is especially important for determining the absolute magnitudes of O stars. Some of the most recent studies of this sort are those by Lesh (1968), FitzGerald (1969), Conti and Alschuler (1971), Walborn (1971a, 1972, 1973a), Burnichon (1973, 1975), and Balona and Crampton (1974). We shall discuss the chief results of Conti and of Walborn, who have succeeded in calibrating their systems of spectral classification in terms of visual absolute magnitude.

Walborn (1972, 1973a) has considered 56 O stars, which are his standard stars for absolute magnitude. These stars belong to 17 stellar groups for which the distance modulus has been determined recently by various authors, generally by means of photometry. If the maximum spread between the different determinations is less than 1 mag, Walborn considers that the modulus is well determined, and he adopts the mean value. In the cases in which the spread is greater than 1 mag, he discusses each case individually. The value of V_0 (the apparent visual magnitude of the O star corrected for interstellar extinction) is calculated from *UBV* photometry, using the intrinsic colors, $(B-V)_0$, of FitzGerald (1970) to obtain the $(B-V)$-color excess, $E(B-V)$, and by adopting, unless otherwise specified, the value $R = 3.0$ for the ratio of the total absorption at the V-band wavelength to $E(B-V)$. The individual values of M_V for the 56 O stars are given in Table V of Walborn (1972) and in Table III of Walborn (1973a).

Table 1-17 lists the clusters and stellar groups used by Walborn, the distance moduli adopted by him, the range spanned by the different estimates of distance modulus for each group, and the number of determinations for each stellar group.

Similarly, Conti has estimated the absolute magnitudes of his O stars from their membership in clusters and associations for which the distances are known independently of the O stars themselves. Conti and Alschuler (1971) have selected carefully the groups for which the distance has been calculated from the spectral types of main-sequence stars later than type O, B-type in general, from a relation giving M_V as a function of spectral type on the zero-age main-sequence (ZAMS) and from *UBV* photometry which is used only to determine $E(B-V)$ and the extinction A_V by means of the formula $A_V = 3E(B-V)$. In some cases, the distance moduli were redetermined by Conti and Alschuler in order to satisfy the above selection conditions. Some of the distance moduli are found also from $uvby,\beta$ photometry. Thus, the calibration of the absolute magnitudes of the O stars rests on the luminosities adopted for the B stars, which luminosities, one hopes, are relatively well known. Conti and Alschuler have discussed in detail each cluster and association used. Conti and Burnichon (1975) added five clusters to this study. A total of 74 O stars classified by Conti and belonging to 19 stellar groups have well-determined absolute magnitudes; the corresponding M_V values for O stars are given by Conti (1975). Table 1-18 lists the clusters and associations used by Conti and the distance moduli which he adopted.

The absolute-magnitude calibrations of Conti and of Walborn are given in Table 1-19. The second column, entitled ZAMS, gives the lower envelope of the M_V spectral-type diagram established by Conti. It is considered by him to be the ZAMS for O stars. Column

Table 1-17
Stellar Groups Used by Walborn for Absolute-Magnitude Calibration of O-type Stars

Stellar Group	V_0-M_V (mag)	Range (mag)	Number of Sources
h Per	11.6	0.6	8
χ Per	11.7	0.6	8
IC 1805	11.0	0.7	5
Per OB2	7.8	0.3	3
Ori OB1	8.3	0.5	10
NGC 2244	10.5	1.4	4
NGC 2264	9.4	0.2	5
Trumpler 14	12.7	–	1
Trumpler 16/Collinder 228	12.0	0.2	3
IC2944	11.5	–	1
Sco OB2	6.0	0.4	3
NGC 6204	12.2	–	1
NGC 6231	11.5	0.8	5
HD 156154 Group	12.1	–	1
NGC 6530	10.9	0.3	3
NGC 6871	12.0	–	1
Lac OB1	8.9	0.1	4

3 gives the median M_V for all the O stars within luminosity class V at each spectral subtype. In a review of the 10 known objects of spectral type O3, Walborn (1982) suggests that his calibration for type O3 I should be revised from -7.0 to -6.4.

Walborn and Conti have thus assigned an absolute magnitude to each of their spectral classes. Although the spectral subdivisions are much more numerous with Walborn than with Conti, the two scales agree well. It should be recalled that these two scales have been established mainly from the same stars.

The equivalent width of Hγ has proved to be a useful criterion of luminosity for the B stars (see *B Stars With and Without Emission Lines*, Underhill and Doazan, 1982, Chapter 2). It has been used also for the O stars, but in the O stars lines of He II blend with the Balmer lines of hydrogen. Balona and Crampton (1974) derived typical absolute magnitudes for O stars from their calibration of the strength of Hγ in terms of M_V, a calibration that is useful chiefly for B stars. Their results are given in Table 1-20, together with the deductions of Lesh (1979), who has compared her original calibration of the MK types for O stars in terms of M_V (Lesh, 1968, 1972) and that of Walborn (1972, 1973a) and has produced a system for O stars of luminosity classes III, IV, and V, which she believes removes some of the inconsistencies of the earlier work.

The calibration of the BCD spectrophotometry in terms of M_V for the O stars by Chalonge and Divan (1973) shown in Figure 1-2 is mainly an extrapolation of the results which were obtained for B stars.

Since the spread in visual absolute magnitude for the O stars is at least 3 mag, it is very difficult to attribute an accurate luminosity to an O star without a precise classification of the star into spectral type and luminosity class. It seems certain that the Of stars are among the most luminous O stars (see Figure 1 of Conti and Alschuler, 1971, and Figure 1 of Conti and Burnichon, 1975). On the other hand, the Oe stars are always situated close to the main sequence, and they seem nearly always to belong to luminosity class V.

Quite recently, Conti et al. (1983a), in a study of the numbers and distribution of O stars and of Wolf-Rayet stars, have reexamined the M_V- spectral type relation for these stars. For O-type stars, the new calibration covers 248 stars in 63 clusters coming for a large part from the Humphreys (1978) study of the brightest stars in the Galaxy. The results are in good agreement with the previous calibrations, confirming the trend noticed by Walborn (1982) for the earliest types; we remark also that the late

Table 1-18
Stellar Groups Used by Conti for Absolute-Magnitude Calibration of O-Type Stars

Stellar Group	Distance Modulus (mag)	Stellar Group	Distance Modulus (mag)
h + χ Per	11.8	Gum Nebula	8.3
IC 1805	11.8	Trumpler 14	12.7
IC 1848	11.8	Trumpler 16/Collinder 228	12.1
Per 2	8.0	IC 2944	11.3
Ori 1b	8.3	NGC 6231	11.5
Ori 1c	8.5	NGC 6530	11.0
Gem 1	10.7	Sgr 4	11.7
NGC 2244	10.9	NGC 6871	11.5
NGC 2264	9.5	Lac 1	8.6
NGC 2362	10.65	Cep 3	9.3

Table 1-19
Visual Absolute Magnitudes of O-Type Stars

Sp. Type	Conti (1975)			Walborn (1973a)						
	ZAMS	V	If	V	IV	III	II	Ib	Iab	Ia
O3	-5.7	-6.1*	-6.7	-5.5						-7.0
O4	-5.5	-5.8*	-6.7	-5.5		-6.4:				-7.0
O5	-5.3	-5.5	-6.7	-5.5		-6.4				-7.0
O5.5	-5.1	-5.3	-6.7							
O6	-4.8	-5.1	-6.7	-5.3		-5.6		-6.3:		-7.0
O6.5	-4.5	-4.9	-6.7	-5.3		-5.6		-6.3:		-7.0
O7	-4.2	-4.7	-6.7	-4.8		-5.6	-5.9	-6.3:		-7.0
O7.5	-4.1	-4.6	-6.7	-4.8		-5.6	-5.9	-6.3:		-7.0
O8	-3.9	-4.5	-6.7	-4.4		-5.6	-5.9	-6.2:	-6.5	-7.0
O8.5	-3.8	-4.4	-6.7	-4.4		-5.6	-5.9	-6.2:	-6.5	-7.0
O9	-3.7	-4.3	-6.7	-4.3	-5.0	-5.6	-5.9	-6.2	-6.5	-7.0
O9.5	-3.6	-4.2	-6.7	-4.1	-4.7	-5.3	-5.9	-6.2	-6.5	-7.0
O9.7							-5.9	-6.2	-6.5	-7.0

*Very uncertain.

Table 1-20
Absolute Visual Magnitude Calibrations of the MK Types for O Stars

Sp. Type	Balona and Crampton (1974)			Lesh (1979)		
	V	IV	III	V	IV	III
O3				-5.5		
O4				-5.5	-6.0	-6.4
O5	-5.4			-5.5	-6.0	-6.4
O6	-5.3			-5.3	-5.4	-5.6
O6.5				-4.8	-5.2	-5.6
O7	-5.1			-4.8	-5.2	-5.6
O7.5				-4.8	-5.2	-5.6
O8	-4.8			-4.4	-5.0	-5.6
O8.5				-4.4	-5.0	-5.6
O9	-4.5	-4.8	-5.0	-4.3	-5.0	-5.6
O9.5				-4.1	-4.7	-5.3

O supergiants are less luminous in the new scale (Table 1-21).

B. O Subdwarfs

The presence in the Galaxy of a population of blue and very blue stars, intrinsically brighter than the white dwarfs on the basis of their spectra and motions, but fainter than the main-sequence stars, was demonstrated for the first time by Humason and Zwicky (1947). Spectra of 48 faint blue stars (apparent magnitude between 10 and 15) in the Hyades region (15 stars) and in the region of the north galactic pole (33 stars) revealed a few white dwarfs and a number of stars with early-type spectra which were apparently normal at the low resolution used. If the stars which are in the direction of the galactic pole had the luminosities corresponding to the main sequence, they would be very far from the galactic plane or even at extragalactic distances. In the direction of the Hyades, only nearby stars are free from interstellar reddening, and the faint blue stars observed in this direction by Humason and Zwicky cannot be at large distances. They have small distance moduli, and their position in the Hertzsprung-Russell diagram is thus well below the main sequence. In fact, we know now (see, for instance, Neckel, 1967) that interstellar extinction in the direction of the Hyades occurs at distances less than a few hundred parsecs and that the reddening at 1000 pc is certainly strong enough to exclude the possibility that a star with $m_V \geqslant 13$ could be observed having a blue color unless it is fainter than at least $M_V = +3$ mag; in the case of hot stars, this limit is several magnitudes below the main sequence. The existence of intrinsically faint blue stars is thus well established, and Humason and Zwicky proposed to identify them with the blue stars at the end of the horizontal branch of globular clusters for which the low luminosities are relatively well known from the cluster's distance. If stellar evolution is well understood, Humason and Zwicky stars can correspond only to late stages of evolution following the exhaustion of hydrogen in the core.

Restricting this discussion to the hottest faint blue stars (i.e., to the sdO types as defined in the paragraph on stellar classification), we must say that the horizontal branch extends to the very high sdO temperatures only in rare globular clusters, in contrast to the fact that sdO types are relatively often encountered among faint blue field stars. Moreover, the kinematical properties of the sdO stars are less extreme than those of the globular clusters.

Table 1-21

M_V-Spectral Type Relation for O Stars (Conti et al., 1983a)

Type	M_V	S.E.	No.	Type	M_V	S.E.	No.	Type	M_V	S.E.	No.
O3 V	-5.4	0.5	5					O3 I	-6.2	—	2
O4 V	-5.8	0.6	5					O4 I	-6.4		3
O5 V	-5.2	0.8	7	O5 III	-6.1		8	O5 I	-6.9	—	2
O6 V	-5.1	0.6	11	O6 III	-5.7	0.3	6	O6 I	-6.5	0.6	3
O6.5 V	-5.1	0.8	13					O6.5 I	-6.7	—	1
O7 V	-4.9	0.5	18	O7 III	-5.6	0.6	6	O7 I	-6.5	1.1	7
O7.5 V	-5.0	—	2	O7.5 III	-5.6	0.4	4	O7.5 I	-6.7	—	1
O8 V	-4.6	0.6	19	O8 III	-5.1	0.5	3	O8 I	-6.5	0.9	6
O8.5 V	-4.9	0.9	4	O8.5 III	-5.0	0.5	3	O8.5 I	-5.9	—	2
O9 V	-4.4	0.6	35	O9 III	-5.3	0.7	16	O9 I	-6.1	0.8	8
O9.5 V	-4.0	0.6	17	O9.5 III	-5.1	0.6	9	O9.5 I	-6.0	0.7	22

Thus, independent empirical information on the absolute magnitudes of these stars is highly desirable. Unfortunately, this type of information is very difficult to obtain for O subdwarfs. They do not belong, like ordinary O stars, to stellar groups of relatively well-known distances, and none of them has a measurable parallax. Some information can be obtained, at least statistically, from the proper motions. Quite a few O subdwarfs have composite spectra with a cool component; if the cool spectrum can be correctly interpreted, it may correspond to a star of known absolute magnitude and give relatively precise information on the luminosity of the O subdwarf component. A very small number of O subdwarfs have been found in globular clusters, and their absolute magnitudes are deduced from the known distance of the cluster. Finally, owing to the absence of better data, the distances of O subdwarfs have been deduced from the mean reddening in their direction or from the strengths of interstellar lines. None of these methods is very satisfying, and the *really empirical* knowledge that we have at the present time about the absolute magnitudes of the O subdwarfs is perhaps not much more precise than it was after the pioneer work of Humason and Zwicky (1947). Spectroscopic absolute magnitudes are now obtained by sophisticated methods involving the comparison of high or medium-dispersion spectra to the best available model atmospheres; hypotheses on the stellar masses are also necessary, and they are suggested by stellar evolution theories. These methods are far from being "really empirical."

We shall now review some of the results obtained.

1. Proper Motions. The population type of the O subdwarfs has been inferred from their proper motions. Generally small but not negligible, these proper motions have a statistical meaning. In some cases, they are complemented by radial velocities. Greenstein and Sargent (1974) have analyzed well-determined proper motions and radial velocities for 11 O subdwarfs and found a mean visual absolute magnitude, $\langle M_V \rangle = +2.9 \pm 0.4$. The tangential velocities indicate a mixture of halo and old disk populations. Some apparently bright O subdwarfs have very small proper motions, and the sdO stars may be a rather inhomogeneous group with parent populations either metal-rich or metal-poor, as suggested also by the existence of composite spectra with a cool star having solar-type abundances.

2. O Subdwarfs with Composite Spectra. Among the faint blue stars, the number of composite spectra is relatively high. They are discovered either spectroscopically, when the two components have about the same contribution in the 4600 to 4000 Å spectral range, or by noting a contradiction between a very blue *U–B* color and a late spectral type. Quite a few composite spectra have been discovered during the course of a search for ultraviolet objects. It is the case for the four last stars of Table 1-22.

We may note the large range in spectral types (A to late G) of the cool component. In some cases, there is a large difference between the visual luminosities of the two components, the sdO star being either the brighter (as in HD 283048) or the fainter (as in BD − 3° 5357).

When the magnitude difference between the two components and the absolute magnitude of the cool star can be estimated (and this is often very difficult), the absolute magnitude of the O subdwarf is known. In Table 1-23, we give the results obtained in the case of HD 128220 and HD 113001, for which we have two or more determinations.

3. O Subdwarfs in Globular Clusters. Searle and Rodgers (1966) discovered the first O subdwarf in a globular cluster, ROB 162 (Woolley et al., 1961) in NGC 6397. Spectroscopic observations at 86 Å/mm showed that ROB 162 closely resembles the sdO stars in having relatively broad Balmer lines and He II lines λ 4686 and λ 4542 much stronger than in normal O spectra. The radial velocity of the star has been measured and indicates cluster membership. Only one other star with a well-established sdO

Table 1-22
O Subdwarfs with Composite Spectra

Star	Spectrum*	Spectral Types given by
BD − 11° 162	sdO + G	Greenstein and Eggen (1966)
Feige 80	sdO + A?	Sargent and Searle (1968)
HD 128220	sdO + gGO	Wallerstein et al. (1963)
HD 113001	sdO + F2	Wallerstein and Spinrad (1960)
BD − 3° 5357	sdO + G8 III	Dworetsky et al. (1977)
HD 283048	sdO + G	Laget et al. (1978)
BD + 10° 2357	sdO + A	Berger and Fringant (1980)
BD + 29° 3070	sdO + F	Berger and Fringant (1980)

*The presence of He II lines in the hot component is really established only in the case of the first four stars.

Table 1-23
Absolute Magnitudes of the sdO Component in Two Composite Stars

Star	M_V (sdO)	Reference
Hd 128220[a]	0	Wallerstein and Wolff (1966)
HD 128220[a]	+4	Goy (1977, 1980a)
HD 128220[a]	+0.6[b]	Gruschinske et al. (1983)
HD 113001[c]	+3.6	Wallerstein and Spinrad (1960)
HD 113001[c]	+3.5	Goy (1980a)

[a]The luminosity class and absolute magnitude of the cool component are very difficult to estimate owing to a very rapid rotation.

[b]M_V is deduced from the distance (d = 550 pc), $E(B-V)$ = 0.07, and values of $x_\lambda = I_\lambda(\text{sdO})/I_\lambda(\text{sdO+G})$ given by the authors.

[c]This star is a visual binary, with Δm estimated by Worley (see Wallerstein and Spinrad, 1960). The cool component has almost no rotation.

spectral type has been shown to be a member of a globular cluster on the basis of its position *and* of its radial velocity; this star is vZ 1128 in M3 (Strom and Strom, 1970). The very blue star K 559 in M15 has the radial velocity of the cluster (Zinn, 1974), but only low-resolution spectra have been obtained and the helium lines are not seen; however, the spectral type is probably sdO. Near the center of NGC 6712, Remillard et al. (1980) discovered a very blue star, C 49, with a well-established sdO type but no radial velocity. The absolute magnitudes of these four stars ($M_V = V$ - apparent distance modulus of the cluster) are given in Table 1-24. The references for the apparent distance moduli are indicated in the third column.

All these stars are well above the blue horizontal branch of the globular clusters and belong to the group of UV-bright stars discussed in Zinn et al. (1972). In a few globular clusters the horizontal branch extends to very blue colors which might correspond to sdO temperatures. An example is the star A 19 in M13 (Savedoff, 1956) for which the UBV photometry is given by Sandage (1969): V = 17.06, B-V =

-0.30, and $U-B$ = -1.09. Even if we do not take into consideration the small reddening of M13, these colors indicate an sdO temperature. With the apparent distance modulus of M13 (= 14.42 mag) given by Sandage (1970), A 19 has an absolute magnitude $M_V = +2.6$ and is much fainter than the sdO stars of Table 1-24.

4. O Subdwarfs with Color Excesses. The distribution of interstellar extinction as a function of direction and distance has been studied by different authors (see, for instance, Neckel, 1967; FitzGerald, 1968; Lucke, 1978; Nandy et al. 1978), using Population I stars with known intrinsic colors and absolute magnitudes. In each direction, the color excess $E(B-V)$ is thus a more or less well-known function of the distance. For O subdwarfs, the intrinsic colors and color excesses are relatively well determined, and thus the distances can be found. However, we must emphasize that the distribution of the interstellar absorbing matter is very irregular and that the use of a mean relation between color excess and distance can give only rough estimates of the distances. Havlen (1976) discovered a low galactic latitude O subdwarf, LS 630, and estimated from the small reddening that the absolute magnitude was $M_V = 5.3 \pm 2$ mag. Upper limits to the absolute visual luminosity of four O subdwarfs have been derived by Walker (1981); the four stars were almost unreddened, and were found to be fainter than $M_V = 0$, $+3.2$, $+5.6$, and $+5.5$, respectively. Schonberner and Drilling (1984) determined absolute magnitudes for a set of 12 very hot O subdwarfs in the galactic plane which had measurable color excesses ranging from about 0.05 to 0.25 in $E(B-V)$. The distances derived from estimates of $E(B-V)$/kpc made by different authors were averaged if they were sufficiently consistent, and the absolute magnitudes thus obtained for 10 O subdwarfs range from $M_V = +1.8$ to $+7.0$ mag.

5. O Subdwarfs and Interstellar Lines. Greenstein (1952) demonstrated the very low luminosity of BD + 28° 4211 from the extreme faintness of the interstellar Ca II K line. Equivalent widths of interstellar lines have been used to derive distances and absolute magnitudes by Kudritzki and Simon (1978) (HD 49798, $M_V = -0.2 \pm 1$ mag) and by Simon (1982) (HD 127493, $M_V \geqslant +3.5$ mag).

6. Spectroscopic Absolute Magnitudes. The comparison of the observed energy distribution in the continuum and of equivalent widths or line profiles to model values yields the effective temperature T_{eff} and the surface gravity g of the star. Such analyses have been made for faint blue (FB) stars (see, for instance, Sargent and Searle, 1968; Greenstein and Sargent, 1974), and in the equation,

$$M_V = 2.5 \log g + 10 \log \theta - 2.5 \log M/M_\odot - 5.82 - BC \, , \tag{1-13}$$

Table 1-24
Absolute Magnitudes of O Subdwarfs in Globular Clusters

Cluster			Star		
Name	Apparent Distance Modulus	Ref.*	Name	V	M_V
NGC 6397	12.40	1	ROB 162	13.24	+0.8
M 3	14.83	2	vZ 1128	14.94	+0.1
M 15	15.29	2	K 559	14.80	-0.5
NGC 6712	15.6	3	C 49	16.6	+1.0

*References: 1 - Alcaino and Liller (1980), 2 - Sandage (1970), 3 - Sandage and Smith (1966).

which is derived (Greenstein and Sargent, 1974) from the definition of the effective temperature (θ being $5040/T_{eff}$ and BC the bolometric correction), the only unknown quantity is the mass M of the star. If it can be proved that the masses of the sdO stars lie in the narrow limits which are observed for the globular cluster horizontal branch stars, the absolute magnitudes can be determined. Greenstein and Sargent (1974) have measured the values of θ and $\log g$ for a sample of 24 O subdwarfs, and adopting the same mass $M = 0.66\ M_\odot$ for all the stars, they calculated the absolute magnitude M_V. Even for those cases in which θ and $\log g$ seem to be relatively reliable, the O subdwarfs form in the $\log g$, θ plane an almost vertical sequence at $T_{eff} \sim 40000$ K which spans the entire range between the main-sequence O stars and the white dwarfs. The absolute magnitudes range from -1.4 to +6.8 and the mean value for the more reliable stars is $\langle M_V \rangle = +4.3 \pm 0.5$ mag with $\sigma(M_V) = \pm 1.7$ mag. We have seen above that 11 O subdwarfs of the same sample with well-determined proper motions and radial velocities had a mean absolute magnitude $\langle M_V \rangle = +2.9 \pm 0.4$ mag. The mean spectroscopic magnitudes of these 11 stars is 4.0 ± 0.6 mag, showing that the luminosity from kinematics is not in very good agreement with the spectroscopic luminosity. This fact and the wide range spanned by the spectroscopic luminosities suggest that the O subdwarfs form a rather heterogeneous group.

C. Wolf-Rayet Stars

Since Wolf-Rayet stars are very rare in our Galaxy and few are to be found near the Sun, determining their distances and their extinctions accurately is difficult. These tasks can be attacked more easily for stars in the Large Magellanic Cloud (LMC) than for galactic stars because the distance modulus of the LMC is relatively well known (to within about ± 0.2 mag) and the interstellar extinction is small. The Wolf-Rayet stars of the LMC have been observed by several astronomers since their recognition by Cannon (1924), and no significant spectroscopic differences have been noted between the Wolf-Rayet stars of our Galaxy and those of the LMC. Smith (1968b) found that stars of classes WC6, 7, 8, and 9 are lacking completely from LMC and that stars of class WN6 are rare. We shall assume that the absolute magnitudes found for the Wolf-Rayet stars of the LMC will be valid for the Wolf-Rayet stars of our Galaxy, at least until the contrary is clearly demonstrated.

It is important to be precise about what is meant by the magnitude of a Wolf-Rayet star. What is the range of wavelengths used, and what are the contributions of the emission lines which may be present in the passbands? The absolute magnitude M_v used here refers to the narrowband photometric system introduced by Westerlund (1966) and described by Smith (1968b). This system is presented at the end of this section.

Smith (1968b, 1968c) has given an absolute magnitude calibration based on observations of Wolf-Rayet stars in the LMC for the subtypes found there and on the observations of galactic Wolf-Rayet stars for the other subclasses. The absolute magnitudes for these latter stars are more uncertain than those for the former stars. Later, Crampton (1971) discussed again the absolute magnitudes of the galactic Wolf-Rayet stars, retaining only those stars associated with nebulae for which the distances come from spectroscopic parallaxes of O and B stars or are found from methods based on the kinematics of O and B stars. The luminosities which Crampton finds are generally fainter than those of Smith. At this time, however, there is no reason to suppose that the visual absolute magnitudes of the galactic and the LMC Wolf-Rayet stars differ significantly. This problem can be studied further only when better data are available on the faint Wolf-Rayet stars in the LMC and on the OB stars which give the galactic distances. Smith (1973) has compared the calibrations of 1968 and of 1971 carefully, keeping only the most reliable values in the one or the other of the scales. In Table 1-25, the M_v adopted by Smith (1968b), the values of

Crampton (1971), and the definitive M_v adopted by Smith (1973) are given. For all WN stars, except for type WN6, as well as for WC5 to WC7, Smith has retained her results of 1968, which come from the Wolf-Rayet stars of the LMC. For other stars, the M_v rests on galactic determinations which include for type WC8 the result determined for γ^2 Vel by Conti and Smith (1972). It appears that all the WC stars have practically the same visual absolute magnitude, a value which is comparable to the value adopted for types WN3, WN4, WN5, and WN6. The WN7 and WN8 stars are more luminous than the remainder of the Wolf-Rayet stars.

Subsequent to the discovery of faint early WN stars in the LMC (Azzopardi and Breysacher, 1979; Massey and Conti, 1983a), to the availability of new spectral types (van der Hucht et al., 1981; Conti et al. 1983b), and to the study of the multiplicity and absolute magnitudes of Wolf-Rayet stars in the LMC by Prévot-Burnichon et al. (1981), the reinvestigation of the M_v calibration of Wolf-Rayet stars by Conti et al. (1983a) leads to the results given in Table 1-26. There is now a progression in luminosity among the early WN types, with all the late WN types culminating around $M_v =$ -6. There also seems to exist a regular increase in luminosity from early to late WC types.

1. The Westerlund-Smith Narrowband Photometric System for Wolf-Rayet Stars. A particularly useful narrowband photometric system was introduced by Westerlund (1966) for observing Wolf-Rayet stars. This system has been described by Smith (1968b). The final system is composed of five interference filters, u', u, b, v, and v', each having a width less than or equal to 230 Å. The characteristics of these filters are given in Table 1-27, together with a list, established by Underhill, of the principal emission features which fall in each passband between the wavelengths defined by the 50-percent transmission points. These lines have been found from the intensity tracings of typical Wolf-Rayet spectra given by Underhill (1959, 1967). Further information on typical shapes of emission lines in the b and v' bands can be found from Figures 5 and 2, respectively, of Bappu (1973). The major ions contributing lines to the broad blended emission features have been selected after a review of the most recent lists of spectral lines known in the

Table 1-25
The First M_v - Spectral Type Relations for Wolf-Rayet Stars

	Smith (1968b)		Crampton (1971)		Smith (1973)
Sp. Type	mag	No.*	mag	No.*	mag
WN3	-4.5	2			-4.5
WN4	-3.9	5	-3.7	3	-3.9
WN5	-4.3	2			-4.3
WN6	-5.8	3	-4.8	7	-4.8
WN7	-6.8	4	-6.5	4	-6.8
WN8	-6.2	3			-6.2
WC5	-4.4	5	-3.6	1	-4.4
WC6	-4.4	0	-3.6	2	-4.4
WC7	-4.4	2	-4.4	3	-4.4
WC8	-6.2	1	-5.4	1	-4.8
WC9	-6.2	0			-4.4

*The number of stars for which individual estimates of M_v exist.

Table 1-26
The Mean M_v - Spectral Type Relation for Wolf-Rayet Stars[a] (from Conti et al., 1983a)

Subtype	M_v	Subtype	M_v
WN2	-2.0	WC4	-2.2:[b]
WN3	-3.6	WC5	-3.9
WN4	-4.0	WC6	-4.3[c]
WN5	-4.8	WC7	-4.5[c]
WN6	-6.0	WC8	-4.8
WN7	-6.0	WC8.5	-5.2[c]
WN8	-5.6	WC9	-5.5
WN9	-5.6[c]		

[a] The M_v must not be overinterpreted.
[b] Adapted from the catalog of van der Hucht et al. (1981).
[c] Extrapolated or interpolated from adjacent types.

laboratory. The relative intensities given in Table 1-27 are those in the WN6 star, HD 192163, or in the WC8 star, HD 192103.

Smith (1968b) has compared her narrow-band photometry of early-type stars with the *UBV* photometry of these stars. The v filter has an effective wavelength close to that of V in the Johnson Morgan system, and in the case of stars without emission lines, Smith can assume that $\theta = V$ to within 0.1 mag. A conversion formula $V-v = -0.02 - 0.36\,(b-v)$, is given in Conti and Smith (1972). For Wolf-Rayet stars, emission lines can account for a large part of the energy in the V band, and the V magnitude may be brighter than the v magnitude by up to 0.5 mag for WN stars and to about 1.5 mag for WC stars. It is clear that for WN stars the visual absolute magnitudes M_v are closely equivalent to those for normal absorption-line stars because no significant emission lines occur in the v filter. For the WC stars, the v band contains a moderately strong emission line at 5135 Å due chiefly to C II, but the presence of emission lines probably does not produce a brightening of more than 0.25 mag (see, for instance, Massey, 1984).

VI. MASSES AND RADII

In this section, we shall review the information available from the study of spectroscopic binaries, as well as that which may be inferred from the positions of O stars in the Hertzsprung-Russell (HR) diagram.

A. Normal O Stars

It is possible to determine the masses of stars directly in only three cases: (1) for visual binaries for which one has a visual orbit and for which a trigonometric parallax exists, (2) for visual binaries for which a visual orbit is known and for which the radial velocities of both components have been measured, and (3) for double-lined eclipsing binaries from analysis of the light curve and the radial-velocity changes. Since O stars are too distant for one to be able to determine visual orbits, the only direct method for obtaining masses is by means of double-lined spectroscopic binaries which are eclipsing. (For those systems which are not eclipsing, one obtains only the minimum masses, $M \sin^3 i$, and for those cases in which the secondary spectrum is invisible, one obtains only the mass function, $f(M) = M_2^3 \sin^3 i/(M_1 + M_2)^2$).

On the other hand, the principal method for determining radii directly uses the fractional radii of the two components of an eclipsing system, R_1/a and R_2/a, which may be derived from an analysis of the light curve. When two spectra are visible and measurable, a can be found in linear units, and linear diameters result directly.

There are about 40 systems containing O stars listed in the catalogs of the orbital elements of spectroscopic binaries by Batten (1967), Pedoussaut and Ginestet (1971), Pedoussaut and Carquillat (1973), and Batten et al. (1978), the number being uncertain because the spectral types of some systems vary from O to B0, depending on the source consulted. More recently, Garmany et al. (1980) have given a list of all known O-type binary systems

Table 1-27
Characteristics of the Narrowband System Used by L. F. Smith (1968b) for Wolf-Rayet Stars

Filter Name	Central Wavelength	FWHM (A)	Emission Lines in WN Stars			Emission Lines in WC Stars		
			Strength*	Wavelength	Constituents	Strength*	Wavelength	Constituents
u′	3500	80	vs	3483	N IV	w	3465	N IV, O III
						w	3494	O IV, S III
						w	3506	C III, O IV, O VI
u	3650	100				m	3608	C III, Al III
						w	3634	Si III, S III
						w	3645	Si III
						s	3689	C IV, N IV
b	4270	70	w	4242	N II, S III	m	4258	C III, S III
						m	4267	C II
v	5160	130				w	5124	C II, O VI
						m	5136	C II
						w	5155	C II, Al III
						w	5171	N II, Al III
v′	5500	230	s	5413	He II, N III, O V, O VI	s	5413	He II, C III, O V, O VI
			m	5474	N II, O V	w	5441	O V
			m	5481	N II	s	5470	C IV, O V
			m	5506	N II, O III	w	5501	O III
						w	5594	O III, O VI

*vs = very strong; s = strong; m = medium; w = weak.

resulting from the compilation of the Batten et al. (1978) catalog and from numerous new spectroscopic studies performed by their group. Columns (1) to (6) of Table 1-28 come directly from that review. Among these systems, few have well-resolved double lines attributable to the one or the other component, and very few are also eclipsing systems which have been well studied photometrically. The number of direct determinations of masses and radii for O stars from eclipsing binaries is then extremely small. The critical review by Popper (1980), who considers only individual results of considerable accuracy determined directly from the observational data, retains only three O-type systems: HD 35921 (LY Aur), HD 100213 (TU Mus), and HDE 228854 (V382 Cyg); the corresponding determinations of masses and radii are given in columns (7) to (10) of Table 1-28 with reference 3 in column (12). We also give some other determinations in columns (7) to (10), with the corresponding reference in column (12); we have tried to eliminate results optimistically derived by some authors, which depend on too many assumptions or are based on dubious data.

The catalog values of $M \sin^3 i$ average around 17.5 and 12 $M_\odot$ for the primary and secondary components. The seven determinations of masses average around 28 $M_\odot$ for M_1 and 24 $M_\odot$ for M_2. The three determinations of Popper give M_1 near 25 $M_\odot$ and M_2 ranging from 8.5 to 19 $M_\odot$. The system HD 47129 (Plaskett's star) is certainly very massive since it has $f(M) = 12.40\ M_\odot$, but this result is very difficult to interpret because the second spectrum probably originates in streams of gas around the secondary star. A model by Sahade (1962), which has been adopted by Stothers (1972), results in $M_1 \sim 51\ M_\odot$ and $M_2 \sim 64\ M_\odot$. A recent study by Hutchings and Cowley (1976) gives the limits 58 $M_\odot < M_1 < 100\ M_\odot$ and 64 $M_\odot < M_2 < 90\ M_\odot$. On the other hand, one can say that no normal O star has a mass significantly less than 20 $M_\odot$. The data are too few in number to attempt to see any correlation between mass and spectral subtype; it is worth noting, however, that only relatively late O stars (⩾O6.5) have well-determined masses and radii from binary studies.

The radii in Table 1-28 range from 4 to 19 $R_\odot$ with the determinations by Popper ranging only from 6 to 13 $R_\odot$. The average in both cases is $\sim 10\ R_\odot$.

Another method for estimating the masses of O stars, which is much less direct than the above method, consists of locating the O stars in a theoretical HR diagram in which evolutionary tracks of stars of different masses have been plotted. This method, used by Conti and Burnichon (1975), results in masses between 20 and 120 $M_\odot$. Such results are indicative only of what may be true, not only because of the uncertainties in the observed quantities used but also because of the uncertain theoretical hypotheses that have been made, which include no mass loss. Masses based on this method and radii resulting from his luminosity and effective temperature calibration for O stars have been given by Conti (1975); they are listed in Table 1-29. For the O subtypes in common, they are in fairly good agreement with the results from binaries.

Angular diameters for stars can be measured by interferometric techniques; they can be found also from lunar occultations or by the method utilized by Underhill et al. (1979) for 160 O and B stars. This method depends on comparing observed absolute energies in intermediate-band filters at long wavelengths with the fluxes predicted from model atmospheres. The angular diameters can be transformed into linear radii when the distance of the star is known. From 18 O stars and 24 early O stars, Underhill et al. (1979) and Underhill (1982) find the average linear radii shown in Table 1-30, which are in satisfactory agreement with the preceding results. Direct measurements of angular diameters with a stellar intensity interferometer have been performed by Hanbury Brown et al. (1974) for three O stars only: ζ Ori (O9.5 I), ζ Pup (O4 ef), and ζ Oph (O9 V(e)). These are in the list of Underhill et al. (1979), and the angular diameters found there are similar, to within 10 percent, to the interferometric data.

Table 1-28
Masses* and Radii* from O-Type Spectroscopic Binaries with Orbital Elements

HD or BD Number	Spectral Type	P (days)	$M_1\sin^3 i$	$M_2\sin^3 i$	$f(M)$	M_1	M_2	R_1	R_2	BFM Number	References‡ (col. 7–10)
1337†	O9.5 III	3.5	10.1	12.9		19	23	13.9	8.9	10	1
						18.3	22.5	11.5	9.5	10	2
12323	ON9 V	3.1			0.004					72	
19820†	O9 IV	3.4	18.9	9.2						115	
25638/9†	O9.5 V	2.7			0.382					157	
35921†	O9.5	4.0	21.6	8.1		21.7	8.4	13.0	10.0	230	3
36486†	O9.5 I	5.7			0.605					233	
37041	O9.0 V	21.0			1.53					241	
37043	O8.5 III	29.1	15.9	9.4						242	
47129	O7.5 III(f)	14.4			12.40	58-100	64-90			286	4
48099	O6.5 V	3.1	0.63	0.38							
57060†	O8.5 If	4.4	20	24		19	23	18.6	14.8	301	5
						23	30	12.3		301	6
57061	O9.0 I	154.9			1.69					302	
75759	O9 V	33.3	17.8	14.3						357	
93205	O3 V	6.1	39	15						416	
93206	O9 Ib(f)	20.7			0.20						
QZ Car	O9 III	6.0			10.5						
93403	O6 (f)	15.1	5.2	3.4						417	
100213†	O7-8	1.39	23.5	15.8		23.8	15.8	7.3	6.2	435	3
135240	O9 V	3.9			1.59					534	
149404	O9 Ia	9.8	1.6	2.7							
150136	O5 III(f)	2.7	14.8	8.2							
151564	O9.5 V	4.6			0.102					586	
E326331	O8 V	5.6			0.412					591	
152218	O9	5.4	13.4	10.7						592	
152219	O9.5 IV	4.2			0.689					593	
152248	O7 If	6.0	24.4	22.5						594	
155775	O9.5	7.0									
159176	O7	3.4	10.8	11.4						627	
165052	O7	6.1	2.5	2.2							
166734	O7 If	34.5	28	28							
167771	O8 I(f)	4.0	2.7	2.3							
175514†	O8.0 V	1.6			0.556	28	10.7	9.6	3.5	719	7
191201	O9 III	8.3	13.8	12.9						786	
E228766	O7 f	10.7	34	23						798	
E228854†	O6.5	1.9	37.3	32.7		26.9	19.0	9.2	7.5	800	3
+40°4220†	O7 f	6.6	31	9						817	
199579	O6.5 III	48.6			0.374					837	
206267	O6	3.7	18.3	6.4						867	
209481	O8.5 III	3.1	6.2	2.9						891	
215835†	O5.5	2.1	23.4	19.1						929	

*In solar units
†Eclipsing system from Batten et al. (1978).
‡Columns (2) to (6): data from Garmany et al. (1980); Columns (7) to (10): data from references listed in column (12) as follows:

1. Wood (1963)
2. Hutchings and Hill (1971)
3. Popper (1980)
4. Hutchings and Cowley (1976)
5. Sahade (1959)
6. Hutchings (1977)
7. Vitrichenko (1971)

Column (11): number in Batten et al. (1978).

In summary, one can say that the masses of O stars lie from 18 to 20 $M_\odot$ to an upper limit ($\sim 120\ M_\odot$?) which is somewhat uncertain, or at least strongly dependent on theoretical hypotheses, because no direct determination is available for the earliest types which are believed to be the most massive.

As to radii, direct determinations fall between 5 and 20 $R_\odot$ and average around 10 $R_\odot$, in general agreement with other determinations for subtypes in common. The results in Tables 1-28, 1-29, and 1-30 show satisfactory agreement. No O star appears to have a truly large radius for its photosphere in spite of the large masses attributed to some O stars and of the spectroscopic evidence for strong winds visible in the ultraviolet absorption lines. Underhill found no evidence that spherical geometry need be used when modeling the photospheres of O-type stars.

B. O Subdwarfs

The general belief that the masses of O subdwarfs cannot be much larger than 1 $M_\odot$ and are probably of the order of 0.5 to 0.7 $M_\odot$ is based on internal structure models, but really empirical knowledge about these masses is very meager. Only one O subdwarf is a component of a spectroscopic binary with radial velocity curves available for the two members, and since no eclipses are observed, only lower limits for the masses can be derived. The velocity curves of this spectroscopic binary (HD 128220, sdO + G) have been analyzed by Wallerstein and Wolff (1966), and they indicate that $M_{sdO} = M_G = 4.2\ M_\odot$. These high masses posed a difficult problem for the interpretation of both stars, but since the velocity curve of the G component is poorly determined, it seems possible to reduce the mass to about 2 $M_\odot$ for the sdO component (Wallerstein and Wolff, 1966) and even 1.2 $M_\odot$ (Gruschinske et al., 1983).

Masses can also be deduced from a spectral analysis at a resolution large enough to give reliable values of T_{eff} and g when the absolute magnitude is known from an estimate of the distance, by writing Equation (1-13) in the form:

$$\log M/M_\odot = \log g + 4 \log \theta - 0.4 M_V - 0.4\, BC - 2.33 \quad . \tag{1-14}$$

The O subdwarfs which are proved members of a globular cluster have the best-known distances, but their faint apparent luminosities render the spectral analysis very difficult. The masses implied by the spectral analysis of vZ 1128 by Strom and Strom (1970) have been calculated by these authors and range from 1.2 $M_\odot$ to 10.5 $M_\odot$ when all sources of uncertainties are taken into account. Using the same

Table 1-29
Masses and Radii of O-Type Stars According to Conti (1975)

Spectral Type	$M/M_\odot$ ZAMS	$R/R_\odot$ V	$R/R_\odot$ If	Spectral Type	$M/M_\odot$ ZAMS	$R/R_\odot$ V	$R/R_\odot$ If
O3	120*	14.5	19.1	O7	28	8.7	22.9
O4	90*	13.5	20.0	O7.5	25	8.3	22.9
O5	60*	11.8	20.9	O8	23	8.3	23.4
O5.5	45	11.0	20.9	O8.5	21	7.9	24.6
O6	37	10.2	21.9	O9	19	7.8	24.6
O6.5	30	9.6	21.9	O9.5	18	7.8	24.0

*Quite uncertain.

Table 1-30
Estimates of $R/R_{\odot}$ for O Stars by Underhill et al. (1979) and Underhill (1982)

Conti Sp. Type	V	III	II	If	Ia
O9.5	7.4 (1)*			18.3 (3)*	36.9 (1)*
O9	8.6 (2)			19.8 (1)	
O8.5		10.1 (1)*		23.8 (1)	
O8		9.2 (3)		30.2 (2)	
O6.5	9.5 (1)				
O6			16.5 (1)*		
O5	12.5 (5)	15.9 (4)		16.2 (2)	
O4	11.9 (5)		17.0 (1)	20.3 (1)	
O3	11.7 (4)			19.8 (1)	

*The number of stars is in parentheses.

spectral analysis but a more recent distance modulus (see Table 1-24), bolometric corrections taken from Table 1-16, and Equation (1-14), the following masses are derived:

T_{eff} (K)	log g	$M/M_{\odot}$
31500	4.5	1.5
31500	4.05	0.5
35000	5.2	6.1
35000	4.6	1.5

This shows that, even in the best cases, the uncertainty in the "observed" values of the masses is still very large.

Quite a few field O subdwarfs have apparent magnitudes bright enough to allow a spectral analysis more precise than for the globular cluster stars, but the distances are unknown or very uncertain. We have seen above that Havlen (1976) derived from estimates of the distance an absolute magnitude $M_V = 5.3 \pm 2$ mag for the O subdwarf, LS 630. This star has been analyzed by Hunger et al. (1981), and their results are reproduced in Table 1-13. The masses calculated by Equation (1-14), which correspond to the maximum, mean, and minimum absolute magnitude of Havlen are respectively 0.033, 0.21, and 1.32 in solar units. From a lower limit of the distance to HD 127493 indicated by the Ca II K interstellar line, Simon (1982) found that $M > 0.3\ M_{\odot}$ and $R > 0.20\ R_{\odot}$. Kudritzki and Simon (1978) have made a very careful analysis of HD 49798, taking into account the atmospheric parameters, the intensity of the Ca II K interstellar lines, and an hypothesis on the axial rotation in order to reduce as far as it was possible the range in probable radii and masses. They obtained $R = 1.45 \pm 0.25\ R_{\odot}$ and $M = 1.75 \pm 1 M_{\odot}$.

In spite of many efforts, the empirical masses are still very uncertain and the conclusion is that values around 0.5 to 1.0 $M_{\odot}$ which are deduced from known evolutionary tracks are the more probable ones. However, some "empirical" values seem to point toward higher masses.

With hot effective temperatures and faint intrinsic luminosities, the O subdwarfs can have only small radii. More direct information on the radii is provided by eclipsing binaries, and the photometric observation of the eclipse of the hot subdwarf component in BD -3° 5357 (Dworetsky et al., 1977) by a cooler and larger star with an observable spectrum (G8 III:, $T_{eff} \sim 4700$ K) showed that the ratio of radii $k = R_{sdO}/R_G$ lies in the range $0.01 < k < 0.03$. Because the radius of a G8 subgiant is of the order of 5 to 10 $R_{\odot}$, the radius of the O subdwarf component is of the order of 0.1 $R_{\odot}$.

Since up to now, the case of BD -3° 5357 is the only one for which direct information on

Table 1-31
Radii* Corresponding to Masses and Gravities Typical for O Subdwarfs

	Masses (solar units)			
log g	1.5	1.0	0.5	0.2
4	2.03	1.66	1.17	0.74
5	0.64	0.52	0.37	0.23
6	0.20	0.17	0.12	0.07
7	0.06	0.05	0.04	0.02

*In solar units.

the radii is available, we shall give the radii implied by the masses generally admitted for O subdwarfs. The radius of a star is defined when the mass and the gravity are known:

$$2 \log R/R_\odot = \log M/M_\odot - \log g + \log g_\odot,$$

with $\log g_\odot = 4.44$. Table 1-31 gives the radii corresponding to the range of gravities observed in O subdwarfs and to masses between 1.5 and 0.2 $M_\odot$.

C. Wolf-Rayet Stars

Many Wolf-Rayet stars belong to binary systems in which the other component is an O star, and the well-known double-lined systems provide us with the only means of estimating the masses of WR stars. However, the radial velocities of WR binaries do not behave as well as those of "normal" binaries. Analyzing the radial-velocity and light curves of WR binaries with eclipses in order to find masses does not work well because the various emission lines of the WR spectrum often give different radial-velocity curves. These problems will be discussed later in Part II of this book; we emphasize only that the masses of Wolf-Rayet stars are still quite uncertain.

A recent review of all known galactic WR binaries with orbit solutions, which includes 13 double-lined and 5 single-lined systems, has been made by Massey (1981, 1982b). The principal results of this review for double-lined binaries are given in Table 1-32, together with a few additional more recent data from Niemela (1983). Column (7) in Table 1-32 gives an estimate of the orbital inclination, i, in order to determine the actual masses because only the minimum masses ($M \sin^3 i$) and the mass ratio (M_{WR}/M_O) are available from the double-lined orbit solutions; for the five eclipsing systems, a lower limit on i can be set; for the other systems, Massey assigns a mass to the O-type companion from its position in the HR diagram, using mass-loss evolutionary tracks (column 8); tentative values of i and M_{WR}, which are given in parentheses in columns (7) and (9), are then deduced.

We see in Table 1-32 that the masses of WR components are often higher than 10 $M_\odot$, the classical value formerly admitted (Kuhi, 1973); Massey notes that 40 percent of WR masses are $\geqslant 20\ M_\odot$. The mean value of WR minimum masses is 11 $M_\odot$ and the mean value of WR masses as estimated by Massey is 17 $M_\odot$ (which corresponds quite well to the statistically most probable orbital inclination of 60°) while the mean mass ratio is 0.50. We agree with Massey and Niemela that the masses known for WC stars in binaries are not significantly lower than those for WN stars in binaries. However, it seems that the conclusion by Massey that there do not appear to be any systematic trends of the masses for different subtypes is too pessimistic. It appears from Table 1-32, as noticed by Niemela (1983), that the early WN stars (WNE) in binary systems show the smallest mean values of masses and mass ratios (11.5

Table 1-32
Masses* of Wolf-Rayet Stars in Galactic Double-Lined Binaries (from Massey, 1981, 1982; and Niemela, 1983)

HD Number	Name	Sp	P (days)	$M_{WR} \sin^3 i$	M_{WR}/M_O	Estimated i (deg)	Basis of i	M_{WR}	Ref.†
68273	γ^2 Vel	WC8+O9 I	78.5	17	0.54	<70	No eclipse	>21	(1)
94305		WC6+O6–8	18.8	15	0.47				(2), (3)
97152		WC7+O7 V	7.9	3.6	0.59	(~35)	M_O ~35	(~20)	(1)
152270		WC7+O5	8.9	1.8	0.36	(~25)	M_O ~60	(~20)	(1)
168206	CV Ser	WC8+O8 V	29.7	11	0.48	(~70)	M_O ~25	(~13)	(1)
90657		WN4+O4–6	8.2	8–11	0.52	50	Eclipse	18	(1)
94546		WN4+O	4.9	8	0.34	—	None	⩾8	(1)
				3	0.50				(2)
E 311884		WN6+O5 V	6.3	40	0.84	(~70)	M_O ~60	(~50)	(1)
E 320102		WN3+O7	8.8	1.8	0.33	(~30)	M_O ~35	(~11)	(1)
186943		WN4+O9 V	9.6	9–11	0.52	(~70)	M_O ~25	(~13)	(1)
190918		WN4.5+O9 I	112.8	0.7	0.26	(~25)	M_O ~35	(~9)	(1)
193576	V444 Cyg	WN5+O6	4.2	9.3	0.40	78	Eclipse	10	(1)
						(~55)	M_O ~45	(~17)	(1)
	CX Cep	WN5+O8 V	2.1	5	0.43	⩾50	Eclipse	5–11	(1)
211853	GP Cep	WN6+O	6.7	—	>0.22	⩾50	Eclipse	(~10–25)	(1)
214419	CQ Cep	WN7+O:	1.6	23	1.19	⩾40	Contact eclipse	⩾23	(1)

*In solar units.
† References: (1) Massey (1981, 1982b), (2) Niemela (1983), (3) Niemela et al. (1983).

$M_\odot$ and 0.4), smaller than those for WC binaries (19 $M_\odot$ and 0.5), while the highest values of masses and mass ratios correspond to the WN6–7 systems, which are unfortunately only three in number (mean values 30 $M_\odot$ and 0.75). The data are certainly too scanty to go further than this indication.

It is worth noting with Massey that the average of the combined minimum masses of the O + O binaries (see Table 1-28) and of the WR + O binaries (Table 1-32) both lie around 30 $M_\odot$. Massey has shown also that the correlation of mass ratios in WR binaries with minimum masses of WR components does not indicate that mass ratios may depend on the orbital inclination of the system (as pointed out by Smith, 1968c), but is only a consequence of the range in masses among WR components which is relatively greater than for their O type companions. The aspects concerning the physics and the evolution of massive binaries will be discussed extensively in Part II.

We shall retain in summary that the mean value of mass ratio of WR + O binaries is 0.50 and that the average mass of WR components is around 17 $M_\odot$.

Direct determinations of radii from WR double-lined eclipsing binaries are not normally possible because these systems are complex and do not display total eclipses. However, Kron and Gordon (1950) have given a complete photometric discussion for V444 Cyg, and the model they have constructed probably remains a basic one although its detailed interpretation is difficult. By adopting the spectroscopic elements of Wilson, which are in good agreement with those in Table 1-32, Kron and Gordon proposed a model for the WR component which consists of a small opaque core of radius 2 $R_\odot$ surrounded by a contiguous luminous and semitransparent disk of radius 7 $R_\odot$, this structure being enclosed by a second much larger envelope of radius 16 $R_\odot$; the O component has a radius of 9 $R_\odot$, and the orbital separation is 37 $R_\odot$.

Another "direct" determination has been performed by Hanbury Brown et al. (1970), who have observed the γ^2 Vel system with the intensity interferometer in the continuum at 4430 Å and in the emission line at 4650 Å due to C III and C IV, in order to find the angular diameter of the star and of the region which emits the C III, C IV line. They found a radius of 17 ± 3 $R_\odot$ for the Wolf-Rayet star and of 76 ± 10 $R_\odot$ for the line-emitting region, provided that the distance to γ^2 Vel is 350 ± 50 pc. However, these results are based on the assumption that the WC8 component is the brighter one in the system (with $\Delta V = 1$ mag), as was formerly thought (Smith, 1968b). It is now believed that the O9 I component is the more luminous member by about 1.4 mag (Conti and Smith, 1972, and references therein), and the study of Hanbury Brown et al. should be revised in view of this. One can say qualitatively that the measured angular diameter is that of the O star, and if we assume equal effective temperatures for the two components, the angular diameter of the WR star is 0.53 times that of the O star. The angular diameter of the line-emitting region remaining unchanged, the radii are ~9 $R_\odot$ for the WC8 photosphere and ~76 $R_\odot$ for the line-emitting region. This shows that the line-emitting region is much more extended than that responsible for the continuum, it being about 8 times as large as the photosphere. If, instead of 350 pc, one adopts the distance derived by Conti and Smith (1972) for the association surrounding γ^2 Vel, the radii are ~12 $R_\odot$ and ~100 $R_\odot$ for the WC photosphere and the line-emitting region, respectively.

The radii of the photospheres of Wolf-Rayet stars can be tentatively determined by the method used by Underhill et al. (1979) for 160 O and B stars. This has been done for 10 Wolf-Rayet stars by Underhill (1983), who found that the photospheric radii of five WN5–6 and WC6–8 stars lie in the vicinity of 10 solar radii, while four WN7–8 stars have larger radii which average around 30 $R_\odot$.*

*A critique of the applicability of this approach to WR stars appears in Chapter 3.

VII. CATALOGS

Although it is now quite easy to get extensive information from various data centers throughout the world, we think it is still useful and convenient (and less expensive) to know where to find general information in the literature. We shall attempt to give here a list of works which are related to O, sdO, and WR stars and which have been of current use in preparing Part I; these references are not always found in Part I.

A. Normal O Stars

We list here the most useful general catalogs and a selection of more specialized lists of observational data.

1. General Catalogs. The *Catalogue of Galactic O Stars* in which Cruz-Gonzalez et al. (1974) have attempted to list all the O stars known to belong to the youngest Population Group of the Milky Way contains 664 stars with photometric and spectroscopic data, the distance, the distance perpendicular to the galactic plane, the radial velocity, the radial component of the peculiar velocity, possible multiplicity, and whether the star is situated inside or outside an H II region.

The *New General O-Type Stars Catalogue* (Goy, 1973) collects data on the stars that have been classified as type O at least once. It is regularly updated, and the fourth edition (Goy, 1980b) contains 971 O stars for which, besides photometric and spectroscopic data, information is given about polarization, situation toward H II regions, duplicity, and variability.

Garmany et al. (1982) have compiled a machine-readable catalog of over 750 galactic O-type stars with published photometry, spectral types, and luminosity classes which is probably complete to a distance of ~2.5 kpc. This catalog is available through the Astronomical Data Center at the Goddard Space Flight Center, NASA (Garmany, 1983). Table 1 of Garmany et al. (1982) lists the 424 stars which are within 2.5 kpc, together with their spectral classification, effective temperature, and bolometric magnitude.

2. O Stars in Associations and Clusters. See Humphreys (1978).

3. O Stars in the LMC. Most of them may be found in the lists of Rousseau et al. (1978) and Walborn (1977).

4. O Stars in the SMC. Most of them may be found in the list given by Humphreys (1983).

5. Spectral Classification and Spectra of O Stars

- Walborn (1972, 1973a)
- Conti and Leep (1974) and Conti and Frost (1977)
- *MK Atlas*, Morgan et al. (1978)
- *Atlas of Yellow-Red OB Spectra*, Walborn (1980)

6. Spectrophotometric Scans. See Breger (1976a, 1976b).

7. Photometric Catalogs. See general catalogs and *ubvy* H_β *Photometry for ~100 O Stars* (Morrison, 1975).

8. Ultraviolet Data. The principal lists of ultraviolet data are given in the section *Main Ultraviolet Programs* at the end of each paragraph. See also *IUE Low-Resolution Atlases* (Wu et al., 1983; Heck et al., 1984), and the high-resolution atlas of Walborn et al. (1985).

B. O Subdwarfs

The number of known O subdwarfs is not very large, and no general catalog is available. The O subdwarfs are recognized by spectroscopic analyses, and candidates are chosen in finding lists of faint blue stars near the galactic poles or in front of absorbing clouds, where blue Population I stars have much brighter apparent magnitudes. The finding lists of faint blue stars contain only a small percentage of O subdwarfs, and they are not given here. They can be found in the papers referenced below. In some cases, O subdwarfs have been selected from ultraviolet surveys by their high ultraviolet color temperatures. We give below references to investigations in which the O sub-dwarfs listed have been analyzed at a dispersion better than 200 Å/mm and for which the spectral type and/or the effective temperature, the gravity, and in some cases the chemical composition have been determined:

- Greenstein and Sargent (1974, Table A4) lists 30 O subdwarfs with their effective temperatures. Values of log *g* are also available for most of the stars, but uncertainty is sometimes quite large. Two more O subdwarfs appear in Table A6 of Greenstein and Sargent as members of composite stars.

- Berger and Fringant (1980) list 12 already known O subdwarfs (among which 4 are composite) and 8 newly discovered ones (among which 2 are composite). Only spectral types are determined here.

- Hunger et al. (1981) determined by a non-LTE analysis the effective temperature, the gravity, and helium content for 11 O subdwarfs. The model-fitting procedure is that of the Kiel group as illustrated by Kudritzki and Simon (1978).

- Heber et al. (1984) list the 26 O and OB subdwarfs for which a non-LTE analysis has been made by using the model-fitting procedure of the Kiel group. The effective temperature, gravity, and helium content are given for the 26 stars, and C, N, and Si abundances are given for 6 of them.

- Schönberner and Drilling (1984) list 12 newly discovered O subdwarfs (Drilling, 1983) and give their ultraviolet color temperatures determined from IUE low-resolution spectra.

C. Wolf-Rayet Stars

After a brief description of the general catalog of galactic Wolf-Rayet stars, we give the up-to-date list of WR stars in the nearest galaxies, as well as some more specialized lists of observational data.

1. General Catalogs

- *Galaxy*—The most recent compilation is the *Sixth Catalogue of Galactic Wolf-Rayet Stars, Their Past and Present,* which has been prepared by van der Hucht et al. (1981). Here five earlier catalogs of Wolf-Rayet stars are compared, and the spectral classification of each star is reviewed critically. A finding chart is presented for each star, and information is given about coordinates, magnitudes (the v magnitude of the narrowband photometry of Smith (1968b), the V magnitude, or m_{pg}), and duplicity. A list is also given of the central stars of planetary nebulae having Wolf-Rayet spectra, and a bibliography on galactic Wolf-Rayet stars of Population I from 1867 to 1980 is presented in an Appendix. A list is given of the stars which have been rejected from previous catalogs.

- LMC—See Breysacher (1981).

- SMC—See Azzopardi and Breysacher (1979).

- M33—See Massey and Conti (1983b).

2. Photometric Catalogs

- *UBV* System

  - Pyper (1966), galactic stars.
  - Feitzinger and Isserstedt (1983) LMC.

- *u'ubvv'* or *ubvr* System
  - Westerlund (1966), northern galactic stars.
  - Smith (1968b), LMC and southern galactic stars.
  - Lundström and Stenholm (1979), faint galactic stars.

3. Scanner Spectrophotometry (3300 to 11100 A) and IR Photometry (1.6 to 11.3 μm). See Cohen et al. (1975).

4. Near-Infrared Spectra

- Sivertsen (1981), 40 Å resolution, 4400 to 9800 Å range.
- Vreux et al. (1983), 8 Å resolution, 5750 to 10350 Å range.

5. Ultraviolet Data. See the principal lists of ultraviolet data given in the section *Main Ultraviolet Programs* at the end of each paragraph, and *IUE Low-Resolution Spectra of 15 Wolf-Rayet Stars* (Nussbaumer et al., 1982).

REFERENCES

Abbott, D.C., and Hummer, D.G. 1985, *Astrophys. J.*, **294**, 286.

Alcaino, G., and Liller, W. 1980, *Astron. J.*, **85**, 680.

Auer, L.H., and Mihalas, D. 1972, *Astrophys. J. Supplement*, **24**, 193.

Azzopardi, M., and Breysacher, J. 1979, *Astron. Astrophys.*, **75**, 120.

Balona, L., and Crampton, D. 1974, *Mon. Not. Roy. Astr. Soc.*, **166**, 203.

Bappu, M.K.V. 1973, in *Wolf-Rayet and High-Temperature Stars*, ed. M.K.V. Bappu and J. Sahade (Dordrecht: Reidel), p. 59.

Barbier, D., and Chalonge, D. 1941, *Ann. Astrophys.*, **4**, 30.

Barlow, M.J., and Cohen, M. 1977, *Astrophys. J.*, **213**, 737.

Baschek, B. 1982, in *Ultraviolet Stellar Classification*, ESA SP-182, p. 115.

Baschek, B., Kudritzki, R.P., and Scholz, M. 1980, in *Proc. Second European IUE Conference*, ESA SP-157, p. 315.

Baschek, B., Kudritzki, R.P., Scholz, M., and Simon, K.P. *Astron. Astrophys.*, **108**, 387.

Baschek, B., and Norris, J. 1975, *Astrophys. J.*, **199**, 694.

Batten, A.H. 1967, *Pub. Dom. Astrophys. Obs.*, **13**, 119.

Batten, A.H., Fletcher, J.M., and Mann, P.J. 1978, *Pub. Dom. Astrophys. Obs.*, **15**, 121.

Beals, C.S. 1938, *Trans. I.A.U.*, **6**, 248.

Berger, J., and Fringant, A.-M. 1980, *Astron. Astrophys.*, **85**, 367.

Blackwell, D.E., and Shallis, M.J. 1977, *Mon. Not. Roy. Astr. Soc.*, **180**, 177.

Boggess, A. et al. 1978a, *Nature*, **275**, 372.

Boggess, A. et al. 1978b, *Nature*, **275**, 377.

Bohlin, R.C. 1982, in *Advances in Ultraviolet Astronomy: Four Years of IUE Research*, NASA CP-2238, p. 68.

Bohlin, R.C., and Holm, A.V. 1980, in *International Ultraviolet Explorer (IUE) Newsletter*, (Greenbelt: NASA), No. 10, 37.

Bohlin, R.C., Holm. A.V., Savage, B.D., Snijders, M.A.J., and Sparks, W.M. 1980, *Astron. Astrophys.*, **85**, 1.

Boksenberg, A. et al. 1973, *Mon. Not. Roy. Astr. Soc.*, **163**, 291.

Breger, M. 1976a, *Astrophys. J. Supplement*, **32**, 1.

Breger, M. 1976b, *Astrophys. J. Supplement*, **32**, 7.

Breysacher, J. 1981, *Astron. Astrophys. Supplement*, **43**, 203.

Burnichon, M.-L. 1973, in *Problems of Calibration of Absolute Magnitudes and Temperatures of Stars*, ed. B. Hauk and B.E. Westerlund (Dordrecht: Reidel), p. 126.

Burnichon, M.-L. 1975, *Astron. Astrophys.*, **45**, 383.

Buser, R., and Kurucz, R.L. 1978, *Astron. Astrophys.*, **70**, 555.

Caloi, V., Castellani, V., and Panagia, N. 1982, *Astron. Astrophys.*, **107**, 145.

Cannon, A.J. 1924, *Harvard Bull.*, No. 801, 1.

Carnochan, D.J. 1982, *Mon. Not. Roy. Astr. Soc.*, **201**, 1139.

Castelli, F., Lamers, H.J., Llorente de Andrés, F., and Müller, E.A. 1980, *Astron. Astrophys.*, **91**, 32.

Chalonge, D., and Divan, L. 1952, *Ann. Astrophys.*, **15**, 201.

Chalonge, D., and Divan, L. 1973, *Astron. Astrophys.*, **23**, 69.

Cherepashchuk, A.M., and Khaliullin, Kh.F. 1973, *Soviet Astronomy*, **17**, 330.

Code, A.D., Davis, J., Bless, R.C., and Hanbury Brown, R. 1976, *Astrophys. J.*, **203**, 417.

Code, A.D., Holm, A.V., and Bottemiller, R.L. 1980, *Astrophys. J. Supplement*, **43**, 501.

Code, A.D., Houck, T.E., McNall, J.F., Bless, R.C., and Lillie, C.F. 1970, *Astrophys. J.*, **161**, 377.

Code, A.D., and Meade, M.R. 1979, *Astrophys. J. Supplement*, **39**, 195.

Cohen, M., Barlow, M.J., and Kuhi, L.V. 1975, *Astron. Astrophys.*, **40**, 291.

Conti, P.S. 1973, *Astrophys. J.*, **179**, 181.

Conti, P.S. 1975, in *H II Regions and Related Topics*, ed. T.L. Wilson and D. Downes (Berlin: Springer-Verlag), p. 207.

Conti, P.S., and Alschuler, W.R. 1971, *Astrophys. J.*, **170**, 325.

Conti, P.S., and Burnichon, M.-L. 1975, *Astron. Astrophys.*, **38**, 467.

Conti, P.S., and Frost, S.A. 1977, *Astrophys. J.*, **212**, 728.

Conti, P.S., Garmany, C.D., de Loore, C., and Vanbeveren, D. 1983a, *Astrophys. J.*, **274**, 302.

Conti, P.S., and Leep, E.M. 1974, *Astrophys. J.*, **193**, 113.

Conti, P.S., Leep, E.M., and Perry, D.N. 1983b, *Astrophys. J.*, **268**, 228.

Conti, P.S., Niemela, V.S., and Walborn, N.R. 1979, *Astrophys. J.*, **228**, 206.

Conti, P.S., and Smith, L.F. 1972, *Astrophys. J.*, **172**, 623.

Corso, G.J. 1975, *Bull. Amer. Astr. Soc.*, **7**, 411.

Crampton, D. 1971, *Mon. Not. Roy. Astr. Soc.*, **153**, 303.

Crawford, D.L. 1970, in *The Spiral Structure of Our Galaxy*, ed. W. Becker and G. Contopoulos (Dordrecht: Reidel), p. 283.

Crawford, D.L. 1973, in *Problems of Calibration of Absolute Magnitudes and Temperatures of Stars*, ed. B. Hauck and B.E. Westerlund (Dordrecht: Reidel), p. 93.

Crawford, D.L. 1975a, in *Multicolor Photometry and the Theoretical HR Diagram*, ed. A.G.D. Philip and D.S. Hayes, Dudley Obs. Report No. 9, p. 17.

Crawford, D.L. 1975b, *Pub. Astron. Soc. Pacific*, **87**, 481.

Crawford, D.L. 1978, *Astron. J.*, **83**, 48.

Crawford, D.L., and Barnes, J.V. 1970, *Astron. J.*, **75**, 987.

Cruz-Gonzales, C., Recillas-Cruz, E., Costero, R., Peimbert, M., and Torres-Peimbert, S. 1974, *Rev. Mexicana Astron. Astrof.*, **1**, 211.

Cucchiaro, A. 1982, in *Ultraviolet Stellar Classification*, ESA SP-82, p.1.

Cucchiaro, A., Jaschek, M., Jaschek, C., and Macau-Hercot, D. 1976, *Astron. Astrophys. Supplement*, **26**, 241.

Cucchiaro, A., Jaschek, M., Jaschek, C., and Macau-Hercot, D. 1977, *Astron. Astrophys. Supplement*, **30**, 71.

Cucchiaro, A., Macau-Hercot, D., Jaschek, M., and Jaschek, C. 1978a, *Astron. Astrophys. Supplement*, **33**, 15.

Cucchiaro, A., Macau-Hercot, D., Jaschek, M., and Jaschek, C. 1978b, *Astron, Astrophys. Supplement*, **35**, 75.

Curtis, R.H. 1932, *Handbuch der Astrophysik*, **V**, 1.

de Jager, C., Hoekstra, R., van der Hucht, K.A., Kampermann, T.M., Lamers, H.J., Hammerschlag, A., Werner, W., and Emming, J.G. 1974 *Astrophys. Space Sci.* **26**, 207.

de Jager, C., Lamers, H.J., and van der Hucht, K.A. 1975, *Astrophys. Space Sci.*, **38**, 313.

Divan, L. 1954, *Ann, Astrophys.*, **17**, 456.

Divan, L. 1956, *Ann. Astrophys.*, **19**, 255.

Divan, L. 1963, *Notes et Informations (Obs. Paris)* **XIV**, No. 1, p. 19.

Divan, L. 1965, *Ann. Astrophys.*, **28**, 70.

Divan, L. 1966 in *Spectal Classification and Multicolour Photometry*, ed. K. Lodén, L.O. Lodén, and U. Sinnerstad (London: Academic Press). p. 311.

Divan, L. 1970, *Astron. Astrophys.*, **12**, 76.

Drilling, J.S. 1983, *Astrophys. J.(Letters)*, **270**, L13.

Dufton, P.L., and Hibbert, A. 1981, *Astron. Astrophys.*, **95**, 24.

Dworetsky, M.M., Lanning, H.H., Etzel, P.B., and Patenaude, D.J. 1977, *Mon. Not. Roy. Astr. Soc.*, **181**, 13P.

Feinstein, A. 1964, *Pub. Astron. Soc . Pacific*. **76**, 399.

Feitzinger, J.V., and Isserstedt, J. 1983, *Astron. Astrophys. Supplement*, **51**, 505.

FitzGerald, M.P. 1968, *Astron. J.*, **73**, 983.

FitzGerald, M.P. 1969, *Pub. Astron. Soc. Pacific*, **81**, 71.

FitzGerald, M.P. 1970, *Astron. Astrophys.*, **4**, 234.

Galecki, Z., Graczyk, M., Janaszak, E., Kolas, R., Krelowski, J., and Strobel, A. 1983, *Astron. Astrophys.*, **122**, 207.

Garmany, C.D., Conti, P.S., and Chiosi, C. 1982, *Astrophys. J.*, **263**, 777.

Garmany, C.D., Conti, P.S., and Massey, P. 1980, *Astrophys. J.*, **242**, 1063.

Garmany, C.D., Massey, P., and Conti, P.S. 1984, *Astrophys. J.*, **278**, 233.

Goy, G. 1973, *Astron. Astrophys. Supplement*, **12**, 277.

Goy, G. 1977, *Astron. Astrophys.*, **57**, 449.

Goy, G. 1980a, *Astron. Astrophys.*, **88**, 370.

Goy, G. 1980b, *Astron. Astrophys. Supplement*, **42**, 91.

Greenstein, J.L. 1952, *Pub. Astron. Soc. Pacific*, **64**, 256.

Greenstein, J.L. 1971, *Proc. IAU Symposium 42, White Dwarfs*, ed. W.J. Luyten (Dordrecht: Reidel), p. 46.

Greenstein, J.L., and Eggen, O.J. 1966, in *Vistas in Astronomy*, **8**, ed. A. Beer (Oxford: Pergamon Press), p. 63.

Greenstein, J.L., and Sargent, A.I. 1974, *Astrophys. J. Supplement*, **28**, 157.

Gruschinske, J., Hamann, W.-R., Kudritzki, R.P., Simon, K.P., and Kaufmann, J.P. 1983, *Astron. Astrophys.*, **121**, 85.

Hackwell, J.A., and Gehrz, R.D. 1974, *Astrophys. J.*, **194**, 49.

Hackwell, J.A., Gehrz, R.D., and Smith, J.R. 1974, *Astrophys. J.*, **192**, 383.

Hanbury Brown, R., Davis, J., and Allen, L.R. 1974, *Mon. Not. Roy. Astr. Soc.*, **167**, 121.

Hanbury Brown, R., Davis, J., Herbison-Evans, D., and Allen, L.R. 1970, *Mon. Not. Roy. Astr. Soc.*, **148**, 103.

Havlen, R.J. 1976, *Pub. Astron. Soc. Pacific*, **88**, 685.

Hayes, D.S. 1970, *Astrophys. J.*, **159**, 165.

Hayes, D.S., and Latham, D.W. 1975, *Astrophys. J.*, **197**, 593.

Heber, U., Hunger, K., Kudritzki, R.P., and Simon, K.P. 1984, in *Observational Tests of the Stellar Evolution Theory*, ed. A. Maeder and A. Renzini, p. 215.

Heck, A., Egret, D., Jaschek, M., Jaschek, C. 1984, ESA SP-1052.

Henize, K.G., Wray, J.D., and Parsons, S.B. 1981, *Astron. J.*, **86**, 1658.

Henize, K.G., Wray, J.D., Parsons, S.B., Benedict, G.F., Bruhweiler, F.C., Rybski, P.M., and O'Callaghan, F.G. 1975, *Astrophys. J. (Letters)*, **199**, L119.

Hiltner, W.A., and Schild, R. 1966, *Astrophys. J.*, **143**, 770.

Hjellming, R.M. 1968, *Astrophys. J.*, **154**, 533.

Humason, M.L., and Zwicky, F. 1947, *Astrophys. J.*, **105**, 85.

Humphreys, R.M. 1978, *Astrophys. J. Supplement*, **38**, 309.

Humphreys, R.M. 1983, *Astrophys. J.*, **265**, 176.

Hunger, K., Gruschinske, J., Kudritzki, R.P., and Simon, K.P. 1981, *Astron. Astrophys.*, **95**, 244.

Hutchings, J.B. 1977, *Pub. Astron. Soc. Pacific*, **89**, 668.

Hutchings, J.B., and Cowley, A.P. 1976, *Astrophys. J.*, **206**, 490.

Hutchings, J.B., and Hill, G. 1971, *Astrophys. J.*, **167**, 137.

Jamar, C., Macau-Hercot, D., Monfils, A., Thompson, G.I., Houziaux, L., and Wilson, R. 1976, *Ultraviolet Bright-Star Spectrophotometric Catalogue*, ESA SR-27.

Jaschek, M., and Jaschek, C. 1982, in *Ultraviolet Stellar Classification*, ESA SP-182, p. 9.

Johnson, H.L. 1966, *Ann. Rev. Astron. Astrophys.*, **4**, 193.

Johnson, H.L. 1968, in *Nebulae and Interstellar Matter*, ed. B.M. Middlehurst and L.H. Aller (Chicago: Univ. Chicago Press), p. 167.

Johnson, H.L., and Borgman, J. 1963, *Bull. Astron. Inst. Netherlands*, **17**, 115.

Johnson, H.L., and Morgan, W.W. 1953, *Astrophys. J.*, **117**, 313.

Johnson, H.M. 1978, *Astrophys. J. Supplement*, **36**, 217.

Kienle, H. 1941, *Zeit. f. Astrophys.*, **20**, 239.

Koornneef, J. 1983, *Astron. Astrophys.*, **128**, 84.

Koornneef, J., and Code, A.D. 1981, *Astrophys. J.*, **247**, 860.

Koornneef, J., Meade, M.R., Wesselius, P.R., Code, A.D., and van Duinen, R.J. 1982, *Astron. Astrophys. Supplement*, **47**, 341.

Kron, G.E., and Gordon, K.C. 1950, *Astrophys. J.*, **111**, 454.

Kuan, P., and Kuhi, L.V. 1976, *Pub. Astron. Soc. Pacific*, **88**, 128.

Kudritzki, R.P. 1976, *Astron. Astrophys.*, **52**, 11.

Kudritzki, R.P. 1980, *Astron. Astrophys.*, **85**, 174.

Kudritzki, R.P. 1981, in *Proc. ESO Workshop, The Most Massive Stars*, Garching, p. 49.

Kudritzki, R.P., Hunger, K., Gruschinske, J., Simon, K.P., and Kaufmann, J.P. 1980, in *Proc. Second European IUE Conference*, ESA SP-157, p. 307.

Kudritzki, R.P., and Simon, K.P. 1978, *Astron. Astrophys.*, **70**, 653.

Kudritzki, R.P., Simon, K.P., Hamann, W.-R. 1982, *Astron. Astrophys.*, **118**, 245.

Kuhi, L.V. 1966, *Astrophys. J.*, **143**, 753.

Kuhi, L.V. 1973, in *Wolf-Rayet and High-Temperature Stars*, ed. M.K.V. Bappu and J. Sahade (Dordrecht: Reidel), p. 205.

Kurucz, R.L. 1979, *Astrophys. J. Supplement*, **40**, 1.

Kurucz, R.L., Peytremann, E., and Avrett, E.A. 1974, *Line-Blanketed Model Atmospheres for Early-Type Stars* (Washington: Smithsonian Inst.).

Laget, M., Vuillemin, A., Parsons, S.B., Henize, K.G., and Wray, J.D. 1978, *Astrophys. J.*, **219** , 165.

Lamers, H.J., Faraggiana, R., and Burger, M. 1979, *Astron. Astrophys. Supplement*, **38**, 227.

Lamers, H.J., Hoekstra, R., van der Hucht, K.A., Kampermann, T.M. 1981, ESA SP-1038.

Lamers, H.J., and Snijders, M.A.J. 1975, *Astron. Astrophys.*, **41**, 259.

Leep, E.M. 1982, in *Wolf-Rayet Stars: Observations, Physics, Evolution*, ed. C.W.H. de Loore and A.J. Willis (Dordrecht: Reidel), p. 41.

Lesh, J.R. 1968, *Astrophys. J. Supplement*, **17**, 371.

Lesh, J.R. 1972, *Astron. Astrophys. Supplement*, **5**, 129.

Lesh, J.R. 1979, in *Spectral Classification of the Future*, ed. M.F. McCarthy, A.G.D. Philip, and C.V. Coyne (Citta del Vaticano: Specola Vaticana), p. 81.

Llorente de Andrés, F., Morales, C., Ruiz del Arbol, J.A., and Pérez Mollá, J. 1981, *Astron. Astrophys.*, **100**, 138.

Lucke, P.B. 1978, *Astron. Astrophys.*, **64**, 367.

Lundström, I., and Stenholm, B. 1979, *Astron. Astrophys. Supplement*, **35**, 303.

Macau-Hercot, D., Jamar, C., Monfils, A., Thompson, G.I., Houziaux, L., and Wilson, R. 1978, *Supplement to the Ultraviolet Bright-Star Spectrophotometric Catalogue*, ESA SR-28.

Massey, P. 1981, *Astrophys. J.*, **246**, 153.

Massey, P. 1982a, in *Wolf-Rayet Stars: Observations, Physics, Evolution*, ed. C.W.H. de Loore and A.J. Willis (Dordrecht: Reidel), p. 121.

Massey, P. 1982b, in *Wolf-Rayet Stars: Observations, Physics, Evolution*, ed. C.W.H. de Loore and A.J. Willis (Dordrecht: Reidel), p. 251.

Massey, P. 1984, *Astrophys. J.*, **281**, 789.

Massey, P., and Conti, P.S. 1983a, *Astrophys. J.*, **264**, 126.

Massey, P., and Conti, P.S. 1983b, *Astrophys. J.*, **273**, 576.

Meade, M.R. and Code, A.D. 1980, *Astrophys. J. Supplement*, **42**, 283.

Mehltreter, J.P. 1961, *Ann. Astrophys.*, **24**, 40.

Méndez, R.H., Kudritzki, R.P., Gruschinske, J., and Simon, K.P. 1981, *Astron. Astrophys.*, **101**, 323.

Mihalas, D. 1972a, *Astrophys. J.*, **176**, 139.

Mihalas, D. 1972b, *Non-LTE Model Atmospheres for B and O Stars*, NCAR-TN/STR-76.

Mihalas, D., and Hummer, D.G. 1974a, *Astrophys. J. Supplement*, **28**, 343.

Mihalas, D., and Hummer, D.G. 1974b, *Astrophys. J. (Letters)*, **189**, L39.

Moffat, A.F.J., and Seggewiss, W. 1978, *Astron. Astrophys.*, **70**, 69.

Morgan, W.W., Abt, H.A., and Tapscott, J.W. 1978, *Revised MK Spectral Atlas for Stars Earlier than the Sun* (Yerkes Observatory and Kitt Peak National Observatory).

Morgan, W.W., Harris, D.L., and Johnson, H.L. 1953, *Astrophys. J.*, **118**, 92.

Morrison, N.D. 1975, *Astrophys. J.*, **200**, 113.

Morton, D.C. 1969, *Astrophys. J.*, **158**, 629.

Morton, D.C. 1970, *Astrophys. J.*, **160**, 215.

Morton, D.C. 1973, in *Wolf-Rayet and High Temperature Stars*, ed. M.K.V. Bappu and J. Sahade (Dordrecht: Reidel), p. 54.

Nandy, K., Morgan, D.H., Willis, A.J., Wilson, R., and Gondhalekar, P.M. 1981, *Mon. Not. Roy. Astr. Soc.*, **196**, 955.

Nandy, K., Thompson, G.I., Carnochan, D.J., and Wilson, R. 1978, *Mon. Not. Roy. Astr. Soc.*, **184**, 733.

Nandy, K., Thompson, G.I., Jamar, C., Monfils, A., and Wilson, R. 1975, *Astron. Astrophys.*, **44**, 195.

Nandy, K., Thompson, G.I., Jamar, C., Monfils, A., and Wilson, R. 1976, *Astron. Astrophys.*, **51**, 63.

Neckel, T. 1967, *Landessternwarte Heidelberg-Königstuhl Veröffentlichungen*, **19**, 1.

Newell, E.B., Rodgers, A.W., and Searle, L. 1969, *Astrophys. J.*, **156**, 597.

Niemela, V.S. 1983, in *Workshop on Wolf-Rayet Stars: Progenitors of Supernovae?*, ed. M.C. Lortet and A. Pitault (Paris: Meudon Obs.), p. III-3.

Niemela, V.S., Méndez, R.H., and Moffat, A.F.J. 1983, *Astrophys. J.*, **272**, 190.

Nussbaumer, H., Schmutz, W., Smith, L.J., and Willis, A.J. 1982, *Astron. Astrophys. Supplement*, **47**, 257.

O'Callaghan, F.G., Henize, K.G., and Wray, J.D. 1977, *Appl. Opt.*, **16**, 973.

Panek, R.J., and Savage, B.D. 1976, *Astrophys. J.*, **206**, 167.

Pearce, J.A. 1930, *Pub, Dom Astrophys. Obs.*, **5**, 110.

Pédoussaut, A., and Ginestet, N. 1971, *Astron. Astrophys. Supplement*, **4**, 253.

Pédoussaut, A., and Carquillat, J.-M. 1973, *Astron. Astrophys. Supplement*, **10**, 105.

Popper, D.M. 1980, *Ann. Rev. Astron. Astrophys.*, **18**, 115.

Pottasch, S.R., Wesselius, P.R., and van Duinen, R.J. 1979, *Astron. Astrophys.*, **77**, 189.

Prévot-Burnichon, M.-L., Lequeux, J., Maurice, E., Prévot, L., and Rocca-Volmerange, B. 1984, *Astron. Astrophys.*, **132**, 389.

Prévot-Burnichon, M.-L., Prévot, L., Rebeirot, E., Rousseau, J., and Martin, N. 1981, *Astron. Astrophys.*, **103**, 83.

Pyper, D.M. 1966, *Astrophys. J.*, **144**, 13.

Relyea, L.J., and Kurucz, R.L. 1978, *Astrophys. J. Supplement*, **37**, 45.

Remillard, R.A., Canizares, C.R., and McClintock, J.E. 1980, *Astrophys. J.*, **240**, 109.

Rogerson, J.B., Spitzer, L., Drake, J.F., Dressler, K., Jenkins, E.B., Morton, D.C., and York, D.G. 1973, *Astrophys. J.*, **181**, L97.

Rossi, L. 1979, *Astron. Astrophys.*, **74**, 195.

Rousseau, J., Martin, N., Prévot, L., Rebeirot, E., Robin, A., and Brunet, J.P. 1978, *Astron. Astrophys. Supplement*, **31**, 243.

Sahade, J. 1959, *Pub. Astron. Soc Pacific*, **71**, 151.

Sahade, J. 1962, in *Symposium on Stellar Evolution* (La Plata, Nov. 1960), p. 185.

Sandage, A.R. 1969, *Astrophys. J.*, **157**, 515.

Sandage, A.R. 1970, *Astrophys. J.*, **162**, 841.

Sandage, A.R., and Smith, L.L. 1966, *Astrophys. J.*, **144**, 886.

Sargent, W.L.W., and Searle, L. 1968, *Astrophys. J.*, **152**, 443.

Savage, B.D., and Jenkins, E.B. 1972, *Astrophys. J.*, **172**, 491.

Savage, B.D., and Mathis, J.S. 1979, *Ann. Rev. Astron. Astrophys.*, **17**, 73.

Savedoff, M.P. 1956, *Astron. J.*, **61**, 254.

Schmidt-Kaler, T. 1965, in *Landolt-Börnstein Tables*, Group VI, ed. H.H. Voigt (Berlin: Springer-Verlag), p. 284.

Schönberner, D., and Drilling, J.S. 1984, *Astrophys. J.*, **278**, 702.

Schultz, G.V., and Wiemer, W. 1975, *Astron. Astrophys.*, **43**, 133.

Searle, L., and Rodgers, A.W. 1966, *Astrophys. J.*, **143**, 809.

Simon, K.P. 1982, *Astron. Astrophys.*, **107**, 313.

Simon, K.P. Jonas, G., Kudritzki, R.P., and Rahe, J. 1983, *Astron. Astrophys.*, **125**, 34.

Sivertsen, S. 1981, *Astron. Astrophys. Supplement*, **43**, 221.

Smith, L.F. 1968a, *Mon. Not. Roy. Astr. Soc.*, **138**, 109.

Smith, L.F. 1968b, *Mon. Not. Roy. Astr. Soc.*, **140**, 409.

Smith, L.F. 1968c, in *Wolf-Rayet Stars*, ed. K.B. Gebbie and R.N. Thomas, Nat. Bur. Stds. Spec. Pub. 307, p. 21.

Smith, L.F. 1973, in *Wolf-Rayet and High-Temperature Stars*, ed. M.K.V. Bappu and J. Sahade (Dordrecht: Reidel), p. 15.

Smith. L.J., and Willis, A.J. 1983, in *Workshop on Wolf-Rayet Stars: Progenitors of Supernovae?*, ed. M.C. Lortet and A. Pitault (Paris: Meudon Obs.), p. III-79.

Snow, T.P., and Jenkins, E.B. 1977, *Astrophys. J. Supplement*, **33**, 269.

Stebbins, J., and Whitford, A.E. 1945, *Astrophys. J.* , **102**, 318.

Stothers, R. 1972, *Astrophys. J.*, **175**, 431.

Strom, S.E., and Strom, K.M. 1970, *Astrophys. J.*, **159**, 195.

Strömgren, B. 1966, *Ann. Rev. Astron. Astrophys.*, **4**, 433.

Thomas, R.N. 1968, in *Wolf-Rayet Stars*, ed. K.B. Gebbie and R.N. Thomas, Nat. Bur. Stds. Spec. Pub. 307, p. 2.

Thompson, G.I., Nandy, K., Jamar, C., Monfils, A., Houziaux, L., Carnochan, D.J., and Wilson, R. 1978, *Catalogue of Stellar Ultraviolet Fluxes* (U.K.: The Science Research Council).

Underhill, A.B. 1958, *Mém. Soc. Roy. Sci Liège, 4th Ser.*, **XX**, 91.

Underhill, A.B. 1959, *Pub. Dom. Astrophys. Obs.*, **11**, 209.

Underhill, A.B. 1966, *The Early Type Stars* (Dordrecht: Reidel).

Underhill, A.B. 1967, *Bull. Astron. Inst. Netherlands*, **19**, 173.

Underhill, A.B. 1980, *Astrophys. J.*, **239**, 220.

Underhill, A.B. 1981, *Astrophys. J.*, **244**, 963.

Underhill, A.B. 1982, *Astrophys. J.*, **263**, 741.

Underhill, A.B. 1983, *Astrophys. J.*, **266**, 718.

Underhill, A.B., Divan, L., Prévot-Burnichon, M.-L., and Doazan, V. 1979, *Mon. Not. Roy. Astr. Soc.*, **189**, 601; microfiche *MN* **189**/1.

Underhill, A.B., and Doazan, V. 1982, *B Stars With and Without Emission Lines*, NASA SP-456.

Underhill, A.B., Leckrone, D.S., and West, D.K. 1972, *Astrophys. J.*, **171**, 63.

Vanbeveren, D., and Conti, P.S. 1980, *Astron. Astrophys.*, **88**, 230.

van der Hucht, K.A., Cassinelli, J.P., Wesselius, P.R., and Wu, C.C. 1979, *Astron. Astrophys. Supplement*, **38**, 279.

van der Hucht, K.A., Conti, P.S., Lundström, I., and Stenholm, B. 1981, *Space Sci. Rev.*, **28**, 227.

van Duinen, R.J., Aalders, J.W.G., Wesselius, P.R., Wildeman, K.J., Wu C.C., Luinge, W., and Snel, D. 1975, *Astron. Astrophys.*, **39**, 159.

Vitrichenko, E.A. 1971, *Pub. Crim. Astr. Obs.*, **43**, 76.

Vreux, J.M., Dennefeld, M., and Andrillat, Y. 1983, *Astron. Astrophys. Supplement*, **54**, 437.

Walborn, N.R. 1971a, *Astrophys. J. Supplement*, **23**, 257.

Walborn, N.R. 1971b, *Astrophys. J. (Letters)*, **167**, L31.

Walborn, N.R. 1971c, *Astrophys. J. (Letters)*, **164**, L67.

Walborn, N.R. 1972, *Astron. J.*, **77**, 312.

Walborn, N.R. 1973a, *Astron. J.*, **78**, 1067.

Walborn, N.R. 1973b, *Astrophys. J. (Letters)*, **180**, L35.

Walborn, N.R. 1974, *Astrophys. J.*, **189**, 269.

Walborn, N.R. 1976, *Astrophys. J.*, **205**, 419.

Walborn, N.R. 1977, *Astrophys. J.*, **215**, 53.

Walborn, N.R. 1980, *Astrophys. J. Supplement*, **44**, 535.

Walborn, N.R. 1982, *Astrophys. J. (Letters)*, **254**, L15.

Walborn, N.R., Nichols-Bohlin, J., and Panek, R.J. 1985, *International Ultraviolet Explorer Atlas of O-Type Spectra from 1200 to 1900 Å*, NASA RP-1155.

Walker, A.R. 1981, *Mon. Not. Roy. Astr. Soc.*, **197**, 241.

Wallerstein, G., and Spinrad, H. 1960, *Pub. Astron. Soc. Pacific*, **72**, 486.

Wallerstein, G., Sturch, C., and Klemola, A.R. 1963, *Pub. Astron. Soc. Pacific*, **75**, 61.

Wallerstein, G. and Wolff, S.C. 1966, *Pub. Astron. Soc. Pacific*, **78**, 390.

Wampler, E.J. 1961, *Astrophys. J.*, **134**, 861.

Wesemael, F., Winget, D.E., Cabot, W., Van Horn, H.M., and Fontaine, G. 1982, *Astrophys. J.*, **254**, 221.

Wesselius, P.R. 1977, in *Proc. IAU Symp. 80, The HR Diagram*, ed. A.G. Davis Philip and D.S. Hayes (Dordrecht: Reidel), p. 125.

Wesselius, P.R., van Duinen, R.J., Aalders, J.W.G., and Kester, D. 1980a, *Astron. Astrophys.*, **85**, 221.

Wesselius, P.R., van Duinen, R.J. and de Jonge, A.R.W. 1980b, *A Catalogue of ANS Observations* (The Netherlands: Space Research Dept., Univ. of Groningen).

Westerlund, B.E. 1966, *Astrophys. J.*, **145**, 724.

Whitford, A.E. 1958, *Astron. J.*, **63**, 201.

Wiemer, W. 1974, *Veroffentl. Astron. Inst. Bonn*, No. 88.

Wilson, R. 1955, *Observatory*, **75**, 222.

Wilson, R. 1957, *Pub. Roy. Obs. Edinburgh*, **2**, No. 3.

Wilson, R. 1958, *Mem. Soc. Roy. Sci. Liège, 4th Ser.*, **XX**, 85.

Willis, A.J., and Wilson, R. 1978, *Mon. Not. Roy. Astr. Soc.*, **182**, 559.

Wolf, C.J.E., and Rayet, G. 1867, *Compt. Rend. Acad. Sci. Paris*, **65**, 292.

Wood, F.B. 1963, in *Basic Astronomical Data*, ed. K.Aa. Strand (Chicago: Univ. Chicago Press), p. 369.

Woolley, R.v.d.R., Alexander, J.B., Mather, L., and Epps, E. 1961, *Roy. Obs. Bull.*, No. 43.

Wray, J.D., and Corso, G.J. 1972, *Astrophys. J.*, **172**, 577.

Wu, C.C., Faber, S.M., Gallagher, J.S., Peck, M., and Tinsley, B.M. 1980, *Astrophys. J.*, **237**, 290.

Wu, C.C. et al. 1983, in *International Ultraviolet Explorer (IUE) Newsletter* (Greenbelt: NASA), No. 22.

Zinn, R. 1974, *Astrophys. J.*, **193**, 593.

Zinn, R.J., Newell, E.B., and Gibson, J.B. 1972, *Astron. Astrophys.*, **18**, 390.

3-89

PART II

One Perspective on O, Of, and Wolf-Rayet Stars, Emphasizing Winds and Mass Loss, With Remarks on Environments and Evolution

Edited by

Peter S. Conti

2

OVERVIEW OF O, Of, AND WOLF-RAYET POPULATIONS

I. INTRODUCTION

These very luminous stars make up only a tiny part of the stellar population of our Galaxy ($\sim 10^{-7}$), yet their luminosity and winds contribute an appreciable radiative and kinetic energy input to their surrounding environments. Furthermore, their presence serves to define the spiral-arm characteristics of our Galaxy and others like it. Their vigorous life and later death, presumably as supernovae, not only may initiate associated star formation in nearby gas clouds, but also will enrich the interstellar medium with the products of nuclear reactions in their stellar interiors. The evolution of massive stars almost certainly controls the appearance of the structure of spiral galaxies and the nature of many of the multitudinous irregular-type galaxies. Such stars are generally not now present in elliptical galaxies, but it is generally believed they existed in appreciable numbers when those objects first formed. Our understanding of the nature and evolution of O, Of, and Wolf-Rayet stars will thus help in clarifying the question of how galaxies evolve.

Part II of this volume discusses four related themes: (1) an observational overview of the spectroscopic and extrinsic properties of O and Wolf-Rayet stars; (2) in Chapter 3, the intrinsic parameters of luminosity, effective temperature, mass, and composition of the stars, and a review of their variability; (3) in Chapter 4, the parameters of the stellar winds (an ubiquitous feature of these luminous stars), both their general features and a review of the radiatively driven model; and (4) in Chapter 5, related issues concerning the effects of the stellar radiation and winds on their environments and a summary of the evolutionary interconnection among these objects.

Although this series of NASA monographs is concerned with nonthermal phenomena in stars, I believe it will be instructive here to discuss some of the extrinsic properties of O and Wolf-Rayet stars. The number and distribution of these objects, in our Galaxy and elsewhere, provide astronomical clues to astrophysical problems, particularly the evolutionary relationship between massive O-type stars and Wolf-Rayet objects. It is also useful to consider the rotational velocities and space motions of these stars to demonstrate that they are not anomalous for the Wolf-Rayet objects, even though such possibilities have been suggested in the literature. In addition, the binary frequency of Wolf-Rayet stars is similar to that of O stars and is not a direct cause of their anomalous appearance, as will be discussed in Section VII.

The by-now conventional belief about O, Of, and Wolf-Rayet stars is that the former represent the hydrogen-core-burning phase of luminous-star evolution, and the latter are generally the helium-core-burning descendants of these *very massive stars*. The Of stars represent intermediate phases in the life of very massive stars in which the omnipresent stellar winds are sufficiently strong to show emission in a few of the subordinate optical lines. The evolution of massive stars is greatly affected by mass loss at the surface and by mixing in the stellar interiors, leading in a natural way from hydrogen-rich O stars to helium-rich Wolf-Rayet stars. (Chapter 5 contains more detail and justification on these themes.) Analogous but low-mass Wolf-Rayet objects which are also highly evolved are sometimes seen as central stars of planetary nebulae. Because these will be the subject of another volume of this series, they will not be discussed further here. (Some characteristics appear in Part I of this volume.)

Let me summarize here some of the background literature. Recent IAU-sponsored meetings on the topics in this monograph include Colloquium 59, *Effects of Mass Loss on Stellar Evolution* (Chiosi and Stalio, 1981); IAU Symposium 99, *Wolf-Rayet Stars: Observations, Physics, Evolution* (de Loore and Willis, 1982); IAU Symposium 105, *Observational Tests of Stellar Evolution Theory* (Maeder and Renzini, 1984); IAU Symposium 108, *Structure and Evolution of the Magellanic Clouds* (van den Berg and De Boer, 1984); and IAU Symposium 116, *Luminous Stars and Associations in Galaxies* (de Loore et al., 1986). Among reviews of stellar winds, I should mention the papers of Cassinelli (1979) and Conti and McCray (1980).

I should also note the ESO Workshop on *The Most Massive Stars* (D'Odorico et al., 1981), and the monographs by de Jager (1980) on *The Brightest Stars* and Kitchin (1982) on *Early Emission Line Stars*. The yearly International Ultraviolet Explorer (IUE) meetings have frequently devoted considerable space to hot luminous stars, and proceedings of those conferences are available. Two recent publications, *Effects of Variable Mass Loss on the Local Stellar Environment* (Thomas and Stalio, 1984) and *The Origin of Non-Radiative Heating/Momentum in Hot Stars* (Underhill and Michalitsianos, 1985), should also be considered as special contributions to this literature. The publication of the proceedings of a workshop on *The Connection Between Nonradial Pulsations and Stellar Winds in Massive Stars* (Abbott et al., 1986) should also be noted.

This portion of this monograph is intended to be reasonably complete as to the published literature up to the end of 1985. Of course, it is impossible to discuss every reference in the field of O and Wolf-Rayet stars exhaustively in the space available. In their sixth catalog of Wolf-Rayet stars, van der Hucht et al. (1981) give an appendix of the literature on Wolf-Rayet stars which attempts to be complete to mid-1981. (Such an effort has not been made for O stars.) I have, however, attempted to be reasonably complete on the literature since then. Space is not available here to discuss X-ray binary systems, several of which contain O-type stars. A recent review appears in Joss and Rappaport (1984).

This field of study is expanding very rapidly, and it is not easy to keep completely current in all separate areas. I have utilized the expertise of several of my associates who have been at the Joint Institute for Laboratory Astrophysics (JILA) to supplement my own contributions to this volume. I am particularly indebted to Katy Garmany for the section on observed wind parameters and to Huib Hendrichs for the discussion of the narrow absorption components in stellar winds. In addition, I have been fortunate to be able to include a section on intrinsic variability by Dietrich Baade, from the Space Telescope European Coordinating Facility (STEC) at ESO in Munich, and one on radiatively driven stellar winds by Rolf Kudritzki and his associates, A. Pauldrach and J. Puls, from the observatory of the University of Munich. Dave Abbott (1986) has also written a short review of the current state of the theory of radiatively driven winds from early-type stars. Because it is being published elsewhere, it will not be specifically included here, although it is closely coupled to Chapter 4.

My section of this book should be considered a very personalized view of the current observational situation concerning O, Of, and Wolf-Rayet stars. I have stressed what I think is important and have not necessarily included all contributions. I have deferred to my senior colleague, Anne Underhill, for her alternative explanations of many phenomena concerning these stars, including their spectra and wind parameters, in Part III of this volume. There is substantial disagreement between us about many issues, particularly the effective temperature scale, the mass-loss rates, the composition, and the evolutionary status of the Wolf-Rayet stars, among other things. I look forward to the eventual resolution of these disparate viewpoints. I would like to acknowledge my personal appreciation to Anne for an invitation to spend a Fulbright year in Utrecht during 1969–1970 interacting and discussing hot-star spectra with her. I regret that we have subsequently diverged rather far from each other in our interpretation of these stars.

Since much of this work was written in 1985, I have only included discussion of a few papers published in 1986. Several of these have had a major impact on our recent thinking about these hot stars, particularly the effective temperature scale. I apologize to all my friends and colleagues for those references inadvertently missed and what might be regarded as a cavalier treatment of some of those mentioned. I hope the overall contribution will be sufficiently complete and critical that it will be of interest not only to the "pundits," but also to the interested astrophysical "bystanders."

I should like to dedicate my portion of this monograph to the memory of Cecilia Payne-Gaposhkin, who found considerable "delight" in the spectroscopic appearance of the enigmatic Wolf-Rayet stars. I will always recall walking and talking with her at the occasion of the XX Liège Colloquium in 1975, in which she exclaimed to me "how much fun" one could have with such stars. It is in the spirit of enjoyment in learning about nature that I dedicate this contribution.

II. CONCEPTUAL FRAMEWORK

I will adopt the conventional point of view that the luminosity of these hot stars gives rise to exceedingly strong radiation pressure, much of which occurs in the far-ultraviolet (FUV) regions. They all have substantial winds, which can for the most part be thought of as spherically symmetric and accelerated (although perhaps not initiated) by the radiative field. Radiation is observed not only from the star itself and its associated wind, but also from the interaction of these physical entities with the surrounding stellar environment of ejected material and the preexisting interstellar medium (ISM). The wind is undoubtedly affected by instabilities in its flow (leading to shocks), and rotation, magnetic fields, and other phenomena in the underlying star, such as nonradial pulsations, may also modify its features.

It would be well to introduce some terms as they will be used in Part II of this monograph. The careful reader will note that not all of these are consistent with usage in other parts of this volume. A *star* is a physical entity of gaseous material held together by the force of gravity, more or less spherical in shape, and deriving its energy from nuclear reactions near its center (or from contraction in the pre-main-sequence phase). A negative gradient of mass density goes outward from the center; this density is not zero at the stellar *surface* (cf. below), but has some finite value continuing outward until it merges with what may be described as the ambient ISM. The merger region is really the outer boundary of the star itself, although the material at this layer is usually thought of as nebular rather than stellar in a spectroscopic sense.

Along with the material density gradient, there is a corresponding energy density gradient. In the star itself, the local temperature also decreases outward and the corresponding wavelength dependence of the radiative energy shifts to larger values. However, there are regions in the star, particularly just above its surface, in which the temperature gradient may reverse or become isothermal; the wavelength dependence of the local radiation will likewise vary. An example of this phenomenon is nonradiative processes and X-ray production in the stellar winds. One speaks of a *stellar interior* and imagines it to be subdivided into a *core* and an *envelope*. The former is used to describe the volume in which all the energy is generated; the latter describes that in which energy is only transmitted outward by convection or radiation. The term "envelope" is often used by those modeling stellar interiors and should not be confused with the use of the word in a stellar atmosphere context.

At some point in the outward migration of energy from the interior of the star, there is a region at which the radiation at typically optical wavelengths can escape freely into space. This volume is commonly referred to as the stellar *atmosphere*. The *surface* or *photosphere* of a star is customarily considered the boundary for the visible continuum radiation. In a classical model *atmosphere*, absorption lines are produced if the temperature is still decreasing outward as it is in nearly all stars. The local region from which radiation escapes depends on the particular transition in question, being deeper for weaker lines. In the Sun and most stars, this visible continuum boundary layer is only a few hundred km thick; therefore, it can be considered small and *plane parallel* with respect to the radius of the star itself. In many O supergiants, the atmosphere may be *extended* and may not be describable by plane parallel geometry. In Be and Oe stars, it may not even be spherically symmetric. Absorption lines in the optical region are not generally observed in Wolf-Rayet stars because no continuing decrease occurs in temperature above the "surface" and the atmosphere is very extended so that emission lines are seen instead.

The mass density continues to decrease above the surface of a star. The substantial nonradiative heating of this higher level material in the Sun and similar stars manifests itself as a *chromosphere* or *corona*. In the hot luminous stars that are the topic of this monograph, this material is moving outward at quite appreciable velocities and has been referred to as a "wind." I prefer to distinguish here between "chromosphere/corona" and "wind" because the terminology in the literature has generally associated the former terms with solar-type stars and the latter with hot luminous ones. This semantic distinction may not be physical (Thomas, 1983), particularly since hot-star winds also appear to be nonradiatively heated.

All hot stars have extensive *P Cygni* features* in resonant transitions in the FUV regions, indicating the ubiquitous presence of a wind. The emission lines and P Cygni features in hot luminous stars come from extended regions which are moving at velocities up to a few thousand km s^{-1}. A relatively dense wind is inferred spectroscopically by the presence of optical emission lines such as Hα and λ4686 He II in Of stars and in the multitudinous emission lines of helium, nitrogen, and carbon ions in Wolf-Rayet stars. A thermal continuum in the infrared (IR) and radio regions from free-free processes in the winds of O and Wolf-Rayet stars can also be detected, in contrast to the stellar radiative flux. I will use the term "wind" to describe all material that is flowing out from the surface of the star. I will not use the term "envelope" for this material because of the possible confusion with the phrase used by astrophysicists in calculating stellar evolution models or the terms "mantle" or "sheath" used elsewhere in these volumes because the words have a connotation of a stationary covering, with an implication of an immediate outer boundary. A wind is anything but static; its main features are dynamic processes. Furthermore, the stellar wind ultimately reaches very far from the star to an interface with the ambient ISM.

The terms "circumstellar shell" or "circumstellar material" are sometimes used in the literature to describe phenomena associated with specific ejection events in which a discontinuity in density is observed. The term

*These combine a violet-shifted absorption profile with a more or less red-shifted emission component.

"puffs" has also been used to describe density enhancements in a stellar wind (e.g., Lamers et al., 1982). The adjective "circumstellar" implies material that has been ejected from a star and remains in the immediate vicinity and in association with it. This term should probably not be used to describe the wind in hot stars, which does not remain closely bound but rather flows out to 10^6–10^7 $R_\star$ (Chapter 5). The energy in the wind of hot stars is sufficient to sweep up considerable preexisting interstellar material. Thus, the material is not primarily made up of stellar ejecta, but is rather the swept-up matter surrounding hot stars. Circumstellar material is a term which nicely describes ejecta from objects in which the velocities are much smaller (e.g., Be stars and red supergiants) and which are primarily stellar in origin.

Let us now consider two *environmental* issues. There are two outer boundary conditions for material surrounding hot luminous stars: that due to the radiation and that due to the wind. These interfaces with the ambient ISM are normally a few to tens of parsecs from the exciting O-type or Wolf-Rayet star. For the radiation, there is an ionized hydrogen region inside which the Lyman continuum photons are found (the classic "Strömgren sphere"). Likewise, there is a region in which the stellar wind pressure is in dynamic balance with the ISM. These boundaries do not usually coincide. The expansion time scale for the Strömgren sphere boundary due to the radiation is a few 10^4 years, and that for the wind is a few 10^5 years (Castor et al., 1975b). Thus, the wind is always expanding into an ionized and already modified ISM. I believe that it would be well to identify this medium with another term; *local stellar medium* (LSM) seems to be an attractive choice. This material, which primarily has the composition of the ISM, has been modified by the radiation and wind pressure from its central object and may include a small admixture of material from the star. The total volumes of LSM in the ISM are large, particularly near OB associations (Abbott, 1982b). It is important to note that neither the Strömgren sphere nor the wind pressure boundary is likely to be spherically symmetric, even if the radiation and wind from the star can be approximated in this way. The preexisting ISM is very inhomogeneous in its density and other properties, and the volumes of the LSM will be highly structured.

What radiation properties do we observe spectroscopically from hot luminous stars? Working our way outward from the *surface*, we would see: continuum radiation and absorption lines characteristic of T_{eff} from the photosphere and emission, absorption, and P Cygni features from the wind along with X rays from this region. Farther from the star, we might detect free-free emission from the wind. Direct imaging might show resolved emission lines and continuum radiation (bubbles and rings) at the radiation and wind interfaces. Spectroscopic narrow absorption features from the LSM might be observed, as would narrow absorption-line features from the line of sight to the star through the ISM. *Interstellar* extinction would also probably be observed, but it may also be caused, at least in part, by the LSM.

Concerning the models which describe these phenomena, I will limit myself to those which have been most successful in confrontations with the observations, realizing that these are idealizations of the complex physical situations among hot luminous stars. For O-type stars, the absorption lines and the continuum regions are, for the most part, formed in a relatively narrow portion of the stellar atmosphere in which the outward velocity is small; the optical emission lines in Of stars and the ultraviolet (UV) P Cygni lines in all these objects form farther out in the wind. For reasons of simplicity, these regions have usually been treated separately in the models which have appeared in the literature. Eventually, they will need to be discussed together, and a completely self-consistent physical treatment undertaken. For this volume, I will follow convention and discuss the absorption-line and continuum-forming regions in Chapter 3 separately from the wind in Chapter 4. The reader should keep in mind this artifact of current hot-star models. In Wolf-Rayet stars, the entire atmosphere is in appreciable motion and must be treated as a wind.

For the atmosphere of O-type stars, the classical assumption of local thermodynamic equilibrium (LTE) fails drastically. The Auer and Mihalas (1972) non-LTE plane parallel models are the starting point for all modern work. These radiative equilibrium (RE)/hydrostatic equilibrium (HE) models are used to predict the strengths and profiles of the hydrogen and helium lines and the continua flux. Line-blanketing is not included. Kurucz (1979) has published LTE plane parallel (RE/HE) models for more or less the same temperature and gravity ranges which also include the effects of line-blanketing. These have generally been used in the literature to predict the emergent continua of hot stars. The limitations of these models must be kept fully in mind when they are used, a fact not always adhered to when one is discussing real stars, as we will see later. Extended model atmospheres in non-LTE have been published by Mihalas and Hummer (1974) and Kunasz et al. (1975); although these use RE and HE and are static, they are useful in some applications. Abbott and Hummer (1985) have published a set of models in which the backscattering of photons from the wind has been included as a boundary condition on the atmosphere. Atmospheric models for O stars will be discussed in detail in Chapter 3, Section II.

The wind models which have been most successful in comparison with the observations are those initiated by Castor et al. (1975a), which are radiatively driven by multiple lines found in the FUV and extreme ultraviolet (EUV) regions. Substantial modifications to this initial work have been published by Abbott (1980, 1982a), who calculated the effects of real lines in these hot stars rather than the idealized cases considered by Castor et al. Pauldrach et al. (1986) have calculated an improved radiatively driven wind model by accounting for the finite disk size of the radiation. The main result will be discussed here in Chapter 4. They also discuss the validity of the Sobolev approximation in the Castor et al. theory. Other modifications of the radiatively driven wind model have also been introduced by Lucy and White (1980), who suggest the possibility of self-generated shocks occurring in the outflowing wind to explain the X-ray observations. Cases of rotational significance and magnetic fields have also recently been considered by Friend and McGregor (1984). A substantive critique of all current wind models has been offered by Thomas (1983).

The models of the interstellar environment take their lead from the seminar paper of Castor et al. (1975b), which considered the wind ejection energy in a dynamical context (see also Steigman et al., 1975; and Weaver et al., 1977). These studies considered only a homogeneous ambient medium. The evolution of a wind-bubble "bubble" into a cloudy medium has recently been considered by McKee et al. (1984). These issues are discussed in more detail in Chapter 5.

III. SPECTROSCOPIC PROPERTIES

A. O-Type Stars

O-type stars are characterized by an optical absorption-type spectrum in which the predominant lines are those of the hydrogen Balmer series—He I and He II—and ions from common elements—carbon, nitrogen, oxygen, and silicon. A subset of O-type stars, called Of, have λ4686 He II (and frequently Hα) in emission. These objects are among the hottest and brightest of all O stars, and as Walborn (1971c) has shown, this emission feature can be used as a luminosity discriminant in the hottest stars.

In the near-IR to wavelengths of just over a micrometer, the overall spectrum of an O star is also primarily an absorption one, but in a number of luminous stars, λ10830 He I and λ10124 He II are seen in emission (e.g., Vreux and Andrillat, 1979). Little spectroscopic work has been done on O stars at even longer wavelengths, but continuum measurements there begin to sample the free-free emission from the stellar wind (Chapter 4).

In the FUV from the atmospheric cutoff to the Lyman limit, several resonant transitions of common ions are seen, invariably with P Cygni profiles. Because their abundances in the stellar winds are relatively large, photons at these wavelengths are absorbed and scattered, producing violet-displaced absorption combined with broadened emission lines. The overall appearance of O-type spectra in the FUV is one of strong P Cygni profiles, along with some weaker photospheric absorption features. Complications can arise from contributions of absorption lines due to both ISM and LSM material.

1. Classification Criteria in the Optical. The formal classification schemes for O-type stars have been reviewed and summarized in Chapter I of this volume, so I will not duplicate them here, but add only some supplementary remarks. There have been basically two classification schemes, closely related but not identical. These were first described by Walborn (1971c) and Conti and Alschuler (1971), respectively, and differ primarily in their philosophy (see Walborn, 1979).

Walborn (1971c, 1973b, 1982b, 1982c) adopted standard stars to define the grid of spectral subtypes, thus following the MK method. Stars were put in various "bins" by their differing appearances in the blue region of the spectrum. In O-type stars, emission lines of He II λ4686 and N III λλ 4634,40 are sometimes seen, leading to a luminosity criterion among the O stars with N III emission (i.e., O((f)), O(f) and Of, depending on whether or not λ4686 He II was in absorption, *filled-in*, or in emission). He also adopted the symbol "f^*" to indicate the status of λ4057 N IV emission (the superscript * indicating N IV present and stronger than N III) and the symbol "f^+" (the superscript $^+$ indicating λ4089,4116, Si IV emission). Parentheses and double parentheses were used to describe intermediate cases. Because these superscript symbols occur only among the earliest O types, where the luminosities are, in fact, not very different (Chapter 3, Section I), I have generally not used them in my work. Walborn also uses the symbols "n" and "(n)" to indicate line widths: stars with very broad absorption lines are assigned the former symbol, intermediate widths the (n), and sharp-lined stars no additional letter, although this procedure is not always consistent with spectral type (e.g., see Walborn, 1973b). Ebbets and Conti (1977) found that the measured line widths of O stars showed a continuum of values and did not separate into broad-lined, intermediate, and narrow-lined cases. Hence, I believe the n and (n) symbols, although possibly useful in some contexts, are not physical and add to the clutter of the spectral types.

Walborn (1973b) also introduced Of?P for two stars (HD 108 and HD 148937) with λ4648, 4650 C III emission present, rather an anomaly among O-type stars. These spectra may well indicate shell phenomena, as Walborn suggests, but the use of "?" is confusing and questionable at best. Several stars were also classified Onfp by Walborn, but these were later shown by Conti and Leep (1974) to be related to Oe (and Be) stars. These authors introduced Oef and O(ef) designations, which I believe are preferable, because they indicate the morphological connection.

The approach followed in Conti and Alschuler (1971) was to make quantitative measurements of a few key lines and attempt to relate them to physical quantities of luminosity and temperature from the beginning.* Although this approach has obvious advantages, it also has two disadvantages which were not so obvious at the time: (1) it is much more difficult and time-consuming to *measure* spectral lines (higher than classification dispersions are necessary, plates must be calibrated, etc.); and (2) the question of where to draw a boundary between, for example, 04 V and 05 V stars, is somewhat arbitrary. With quantitative measurements, one would like the observed parameters such as the λ4471 He I/λ4542 He II line

*Mihalas (1984) has pointed out some inherent problems in principle with such an approach.

ratio to change smoothly with spectral type. However, spectral classification necessarily quantizes stars into spectroscopic "bins" in which the properties change discontinuously between adjacent types. Conti and Alschuler adopted quantitative line-strength criteria which agreed with "most" stars of that spectral type. Hence, a few stars were later found to differ with Walborn's classifications. In one case, one of his "standards" had observed properties which were not intermediate with adjacent types. Fortunately, it has turned out that, with only a few exceptions, the Walborn lists (1971c, 1972, 1973b, 1982c) and Conti and Alschuler (1971), Conti and Leep (1974), and Conti and Frost (1977) lists agree fairly well. The discrepant entries probably should be viewed as indicating the uncertainties in any classification scheme.

To summarize then, I would say that classification of most O stars can be safely made to a subtype (e.g., O6 V, O6.5 V, etc.), but in a small (~ 10 percent) fraction of the cases they could be wrong by one subclass. Two subclass errors are probably of the order of 1 percent. Luminosity class distinctions (I, III, V) are probably similarly known. It has been our practice to adopt Walborn types if we have not *measured* plates and found the line-strength ratios (Part I of this volume). We generally do not use the superscripts or "n" notation quoted by Walborn for the reasons discussed above.

2. Classifications in the Yellow-Red. Walborn (1980) has presented an atlas of yellow-red OB spectra for a moderately extensive sample of O3–B9 stars (and a few Wolf-Rayet stars). He essentially shows that the spectroscopic properties observed in this wavelength region are consistent with those which define the system in the blue. Only one hydrogen line, Hα, is available but helium, carbon, nitrogen, and oxygen lines are available and, in some cases, represent ionization states which are not observed at shorter wavelengths. λ5646 C III emission is present in a number of luminous late-type O stars; λ4634, 4640 N III is usually found in emission only in early-type O stars, but overlap occurs among the O6–O8f stars where both features are found to be in emission. Since there are fewer lines in the yellow-red, compared to the blue, it has been appropriate to use this longer wavelength region only as subsidiary spectroscopic information, although Hα has turned out to be a useful diagnostic when found in emission (e.g., Conti and Leep, 1974).

3. Ultraviolet Classification. Among the first systematic surveys of early-type stars was that of Panek and Savage (1976) on OAO-2 (Orbiting Astronomical Observatory). The Copernicus satellite data were analyzed by Bidelman (1977) and Abbott et al. (1982), although they were concerned more with B supergiants than O-type stars. Henize et al. (1981) discussed Skylab observations of a number of OB stars, particularly the resonance Si IV and C IV lines which are very prominent in early-type stars. These P Cygni lines are formed in the stellar wind; resonant absorption lines from the ISM (and LSM) are also strong features in UV spectra.

Bruhweiler et al. (1981) have found a multitude of photospheric Fe V lines below 1500 Å and Fe IV lines longward of this region which are present in many O-type stars. λ1640 He II is seen in absorption in a number of O stars of later types, even some with λ4686 He II in emission or filled-in (Walborn and Panek, 1984b). Nandy (1982) has presented UV spectra of early-type stars based on IUE data.

The most thorough examination of IUE spectra of early-type stars is that of Walborn and Panek (1984a, 1984b) and Walborn et al. (1985). These authors concentrated on the short-wavelength IUE range ($\lambda\lambda$ 1200–1900 Å) where most of the lines are seen; few O-star lines are found between $\lambda\lambda$ 1900–3100 Å in the long-wavelength IUE range. They have presented a detailed examination of ultraviolet classification of many OB stars based on considerable high-dispersion IUE data. A good correlation with optical types is found, thus confirming previous suggestive but less accurate data. They also found that the N V and C IV P Cygni wind profiles change smoothly with spectral type and luminosity class (the absorption profile becoming stronger and the "edge" velocity increasing with earlier types and brighter stars), indicating that these wind parameters are associated with the fundamental parameters of the stars.

The high-dispersion resolution of IUE rebinned to a constant wavelength resolution of 0.25 Å was normalized by Walborn et al. (1985), and an extensive atlas was presented. They illustrate both main-sequence changes and luminosity effects. Figure 2-1 shows an example from Walborn and Panek (1984a). They point out a strong luminosity dependence for the Si IV resonance lines at λλ 1394,1403. These go from pure absorption in main-squence stars to full-blown wind P Cygni profiles in luminous supergiants. The N V and C IV resonance lines show wind phenomena among nearly all O-type stars, being stronger in more luminous stars and covering a more extensive velocity range. A few exceptions to the Si IV phenomenon among intermediate-luminosity O-type stars are noted by Walborn and Panek. Figure 2-2 gives an illustration of the changes along the main-squence OB stars (from Walborn and Panek, 1984b). Here they note the change in the C IV λλ 1548,1550 lines. These authors suggest that features such as these can be used for UV spectral classification, even though they represent wind features rather than photospheric ones.

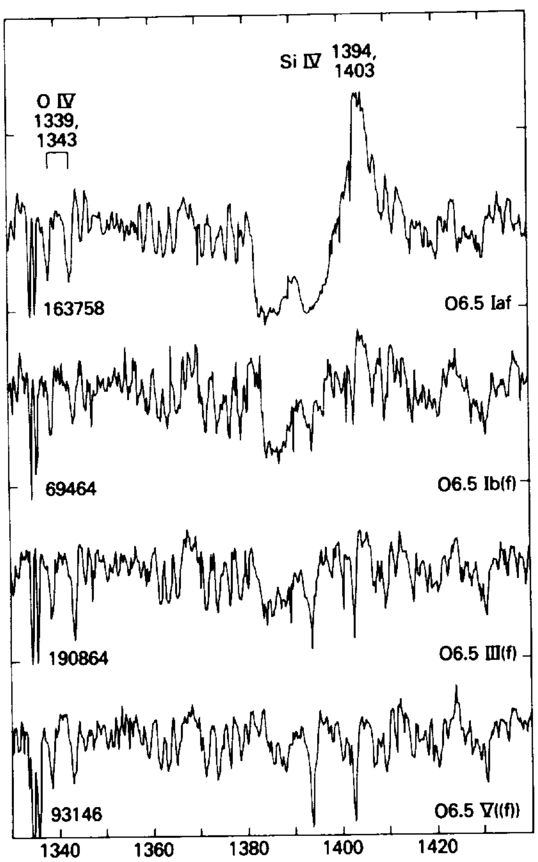


Figure 2-1. Ultraviolet spectral classification (from Walborn and Panek, 1984a). Luminosity changes in λλ 1397,1403 Si IV at spectral type O6.5.

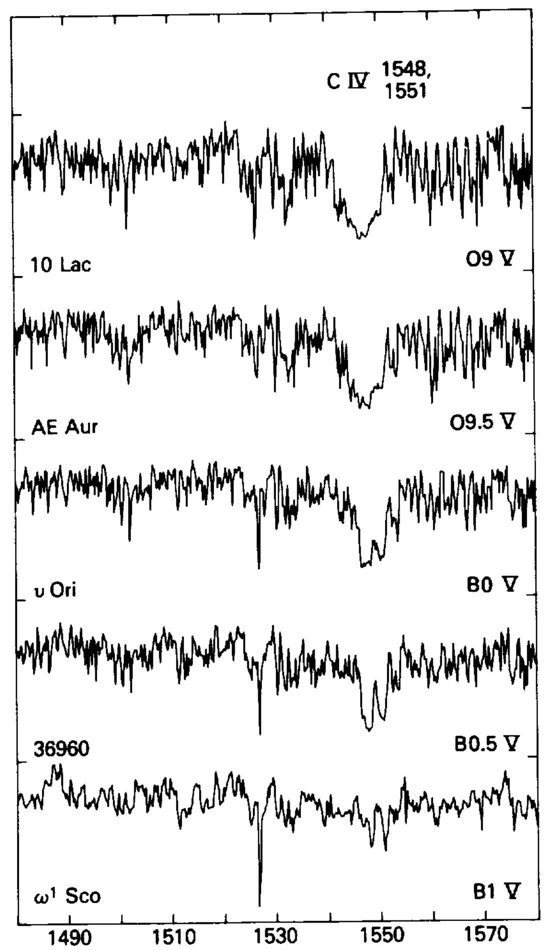


Figure 2-2. Ultraviolet spectral classification (from Walborn and Panek, 1984b). Changes along the OB main sequence in λλ 1548,1550 C IV.

(This issue is still controversial (e.g., Mihalas, 1984).)

4. Line Strengths Among O-type Stars. Photographic plates at coudé dispersion of ~16 to 20 Å/mm can be used to extract reasonably accurate absorption-line equivalent widths. My associates (Conti and Alschuler, 1971; Conti, 1973a, 1973b; Conti and Leep, 1974; Conti and Frost, 1977) and I discussed the strengths of a large number of lines in many O-type stars in our Galaxy. (A few stars had previously been discussed by Peterson and Scholz (1970).) The hydrogen and helium lines were used to estimate the effective temperatures and gravities (Chapter 3, Section IV). Other lines of carbon, nitrogen, oxygen, and silicon ions were also measured, but at the time (this is still mostly true), non-LTE predictions for the strengths of these ions in O-type stars were not available. The measuring uncertainties of the 10 to 20 percent were small enough to enable us to deduce a gross property of the stellar atmospheres such as effective temperature, but were not good enough to determine accurate line profiles, although that was attempted by Conti and Frost.

Looking at the log *W* results for hydrogen and helium lines in those papers, we typically found dispersion at a given spectral type of 0.2; part of this was undoubtedly due to a gravity dependence. (We now know that lower gravity models should have been used for the hotter stars, according to the work of Kudritzki and associates (Chapters 3 and 4).) By contrast, the equivalent width difference of λλ 4648,4650 C III, λλ 4067,4069,4070 C III, λλ 5801,5812 C IV, λ4514 N III, λ3759 O III, λ5592 O III, and λ4089 Si IV were considerably larger at a given spectral type. In most cases, by comparing strengths in dwarfs and supergiants, one could infer that a luminosity (gravity) effect is involved, with the lines being stronger in brighter stars. However, the relevant calculations have not yet been made. Curiously, λλ 4067,4069, 4070 did not exhibit a luminosity dependence, but had a comparable scatter in absorption-line equivalent width. Among a few O-type stars, the λλ 4648,4650 C III and λ4089 Si IV lines were observed in emission; the other lines were not. These ions also tended to weaken with advancing spectral type, presumably mainly an ionization effect, but also one of opacity since C IV was not seen in the hottest main-sequence stars.

The N III λλ 4634,4640,4641 lines are found in emission in many O-type stars. The upper level of this $3d^2D + 3p^2P^0$ transition is populated by dielectronic recombination (from N IV), and the lower level is drained by two electron transitions to $2p^2$ levels according to Mihalas et al. (1972) and Mihalas and Hummer (1973), who have presented the relevant calculations. The $3p^2P^0 - 3s^2S$ transitions (λλ 4097,4103) are found to be in absorption, as is observed. Mihalas et al. point out that these N III lines are in emission as a consequence of atomic physics, rather than the need to invoke an extended atmosphere. The calculated line strengths are consistent with observed ones for O((f)) stars (i.e., those with λλ 4686 He II in absorption). The observed emission lines are stronger in O(f) and Of stars (i.e., those with λ4686 He II filled-in or in emission), as would be found in stars with an extended atmosphere (Auer and Mihalas, 1972; Mihalas and Lockwood, 1972).

O((f)) stars are therefore normal O stars, of generally early subtype, in which N III lines are in emission because of a sufficiently high population of N IV. O(f) and Of stars are more luminous objects with extended atmospheres in which λλ 4686 He II is filled-in or in emission. Of course, all of these stars have winds; in the O(f) and Of stars, the wind is stronger (denser) so that the λ4686 transition occurs in an already finite and extended region. One should be aware, then, that O((f)) and Of stars are not generally identical, and care must be taken in interpreting their evolutionary state.

Unidentified emission lines at λ4485.76 ± 0.04 and λ4504.21 ± 0.07 appear in 24 luminous O stars (Conti, 1973a). These lines were first noted by Wolff (1963) and are still not assigned to a given ion or transition. The shorter wavelength line is about 25 percent stronger, and the two features are clearly related. No absorption lines are present at these wavelengths in other O- or B-type stars. They are found among the O7–O9.5 (and a few B-type) supergiants.

Strong absorption is found for λλ 4603,4618 N V in some very early O-type stars (e.g., Walborn, 1971b). Their spectroscopic properties are somewhat like those of the so-called transition WN stars.

Ebbets and Wolff (1981) have studied the behavior of λλ 9701,9718 C III in a number of early-type stars. They find this transition to be in emission in many O-type stars, being found in stars as late as BO-type supergiants. They also summarize the line strengths and relationships among the C III emission lines frequently observed in emission in O-type stars. Nussbaumer (1971) had calculated the intensity of C III transitions λλ 5696,8500 (singlets) and λλ 4648,4650,9701,9718 (triplets) for a simple (i.e., monotonic velocity law, isothermal conditions) expanding atmosphere. He found that, at the relevant temperatures, λλ 9701–9718 and λ5696 should be in emission, whereas λ8500 and λλ 4648,4650 should be present in absorption. Improved calculations of the singlets were made by Cardona-Nunes (1978) and the triplets by Sakhibullin et al. (1983). Other lines of C III and C IV have been calculated by Sakhibullin and van der Hucht (1982). Ebbets and Wolff find that their measurements of the IR C III triplet in emission are consistent with the expectation that the lower level of the transition is "drained" (via an UV transition), thus underpopulating it. Similarly, the lower level of the singlet, λ5696, is underpopulated by UV transitions and is thus forced into emission. These lines, like the λλ 4634,4640 N III transitions, are in emission because of selective atomic physics, not atmospheric structure, although the latter may strengthen the lines.

5. OBN and OBC stars. Walborn (1970, 1971a, 1976) has introduced and reviewed information concerning OB stars with inordinately strong nitrogen lines and weak carbon lines and with strong carbon lines and correspondingly weak nitrogen lines. The helium and silicon ions provide the reference frame for classification. The lines of carbon, nitrogen, and oxygen are used as additional (mostly luminosity-dependent) criteria. Walborn draws attention to a number of OB stars in which the nitrogen or carbon is strong and/or weak with respect to his classification scheme. These stars also generally stood apart from others in Conti's (1973a, 1973b) measured line analyses. Walborn and Panek (1985) have presented UV spectra of OBN and OBC stars, showing their similar morphology to the optical spectra.

It is a curious fact that the OBN stars are found both near the main sequence (type V) and far from it (type I), yet OBC stars are all very luminous and are well removed from the zero-age main sequence (ZAMS). Very early O types do not show OBN or OBC phenomena. These spectroscopic anomalies may represent real abundance differences: the OBN stars would be stars in which hydrogen-burned nitrogen-rich material is seen; similarly, OBC stars have enhanced carbon (at the expense of nitrogen), and oxygen also seems to be enhanced in these objects. Walborn (1976) discusses possible origins of these anomalies, including initial composition and mixing in certain stars, and postulates that the OBC stars are of normal composition. The more numerous stars classified as normal stars are then nitrogen-rich, and the OBN stars are extreme cases of mixing during the normal main-sequence evolution. These explanations are not entirely satisfying. I would also like to stress that, because extensive abundance analyses have not yet been made, "normalcy" is defined only by the "majority" spectroscopic appearance.

Bisiacchi et al. (1982) have investigated the spectra of 95 O-type stars for the line strength at λ4514 N III. They find a continuum of strengths of this line in the ON stars which show the most extreme values. They therefore suggest that the nitrogen anomaly is a continuous one among all O stars, extending even to the earliest types. López and Walsh (1983) find that

$H\alpha$ emission appears to be a common occurrence among the OBC supergiants, thus suggesting that it is an extended atmosphere effect rather than an abundance effect.

Taken together, these data would suggest that all O stars mix material to the surface to a greater or lesser extent; this matter is rich in CNO equilibrium H-burning residue in which nitrogen is enhanced. The OBN stars are then the extremes of the process. This is an important conclusion if true, but stellar structure models do not as yet support such total mixing. By contrast, all OBC stars appear to be very luminous supergiants, perhaps all having extended atmospheres in which $H\alpha$ is in emission and the C III lines are inordinately enhanced. One would then infer that these stars do not have an anomalous composition of carbon. A continuum of properties would also be expected.

6. Spectra of Metal-Poor O-Type Stars. Crampton (1979) and Crampton and Greasley (1982) obtained spectrograms of a sample of near-main-sequence O-type stars in the Large Magellanic Cloud (LMC) and Small Magellanic Cloud (SMC) in order to determine their distance moduli. As far as the O-type stars were concerned, Crampton did not particularly notice any anomalous line-strength ratios in the LMC stars, compared to galactic standards, at least as could be found on the image-tube plates he was using. Measures of helium and silicon lines compared to Conti's (1973a) galactic stars revealed no differences. Pagel et al. (1978) have found that the gas in H II regions in the LMC is deficient in some elements. Walborn (1977) has noted subtle differences in LMC O9–B0 supergiants from his silicon-helium classification grid (Si IV being marginally weak). Dufour (1984) has summarized the evidence for the lower abundances in the LMC.

On the other hand, for the SMC, the spectroscopic evidence of anomalous line strengths is clear. Crampton and Greasley (1982) found obvious differences in the C III, N III, and Si IV lines in O main-sequence stars, compared to the LMC and galactic stars. Other spectroscopic evidence (Osmer, 1973; Walborn 1977, 1983a; Hutchings, 1982) also indicated "metal" weakness in the SMC, compared to galactic supergiants. Factors of 2 to 3 in the line strengths appear to be consistently measured by different authors. The gas in H II regions is also deficient (e.g., Pagel et al., 1978; Dufour, 1984).

Owing to the general faintness of O-type stars in the Magellanic Clouds, detailed analyses of their spectra have not yet been published. However, improved instrumentation at the European Southern Observatory at La Silla, Chile, has made possible a direct attack on O-type spectra, and Kudritzki and associates are gathering high-resolution data on O-type stars in these nearby galaxies. The results of a planned "fine analysis" will be very helpful in our understanding of O-type photospheres in situations with lower light-element abundances.

B. Wolf-Rayet-Type Stars

The optical spectra of Wolf-Rayet stars are dominated by emission lines of helium, nitrogen, carbon, and oxygen. Beals and Plaskett (1935) suggested two major subgroups: (1) the WN, in which the lines of helium and nitrogen dominate, and (2) the WC, in which helium, carbon, and oxygen ions are seen. In most Wolf-Rayet stars, hydrogen lines are not readily apparent. Underhill (1968) reviewed what might now be referred to as the qualitative situation regarding the optical spectra of Wolf-Rayet stars. Quantitative measurements of emission lines in a few Wolf-Rayet stars were subsequently published (e.g., Bappu and Ganesh, 1968; Smith and Kuhi, 1970; etc.). The classification of Wolf-Rayet stars and determinations of subtypes was put into place by Smith (1968a). In the near-IR regions, Wolf-Rayet stars still show a predominant emission-line spectrum with common ions of HeCNO being observed, depending on the WN or WC nature.

With the following exceptions, absorption-line features are not usually seen in the optical or IR spectra of Wolf-Rayet stars:

- Some Wolf-Rayet stars have O-type companions, either in proven orbit (e.g., V444 Cyg) or in a more distant association (e.g., HD 219460, separation 1″ but not always resolved at the telescope). In these cases, an absorption-line spectrum (most visible as the upper Balmer series) is superimposed on the Wolf-Rayet continuum. The emission spectrum often appears washed out as it is in the usual case where a companion O star is brighter.

- Many Wolf-Rayet stars show P Cygni profiles in selected transitions (e.g., λ4604,4618 N V in WN stars, λ5696 C III in WC stars). These absorption features are formed in the stellar wind.

- A very few WN stars show displaced intrinsic absorption in the upper Balmer series members. This was first demonstrated conclusively by Niemela (1974) in her Ph.D. thesis on HD 92740, in which she showed that the observed Balmer lines moved in phase with the emission features in this orbital binary system; the lines therefore had to be intrinsic to the WN star. In the few WN stars with Balmer lines in absorption, the velocities indicate outflow velocities of approximately 100 km s^{-1} at their formation level (Conti et al., 1979).

- In a few other Wolf-Rayet stars (e.g., HD 193793 (WC6 + abs)), relatively undisplaced absorption lines of the upper Balmer series are observed. These may be intrinsic (Fitzpatrick, 1982) due to an orbiting companion (Lamontagne et al., 1984) or a distance line-of-sight associate (Conti et al., 1984). The true nature of these WR + abs systems (van der Hucht et al., 1981) remains as yet unexplained.

In the FUV regions, emission lines dominate the appearance of Wolf-Rayet stars, with the P Cygni profiles of resonant (and metastable) transitions of C III, C IV, N IV, N V, O VI, and Si IV being very strong. Many subordinate lines of these common ions are seen with pure emission profiles. Similarly to the optical regime, the overall spectra of Wolf-Rayet stars are of their wind features; atmospheric absorption lines are observed in only a very few objects.

Conti et al. (1983b), Perry (1983), Torres and Conti (1984), and Torres (1985) have analyzed numerous optical emission lines in many Wolf-Rayet stars. These studies are based mainly on image-tube spectrograms of 26 Å/mm in the blue and 52 Å/mm in the red obtained at both the Kitt Peak National Observatory and the Cerro Tololo Observatory. The equivalent widths and line profiles measured represent a first reconnaissance of optical spectra for many galactic and LMC WN-type stars. Considerably better line profiles can now be extracted in selected stars with the improved electronic devices becoming generally available on large and moderate-sized telescopes.

Among the first results from this extensive new body of spectroscopic data was that absorption lines due to O-type companions were not an ubiquitous feature among Wolf-Rayet stars (Vanbeveren and Conti, 1980), and thus not all Wolf-Rayet stars were binaries, contrary to a view that had become popular over the years (e.g., separate remarks by Conti, Kuhi, and Sahade in Bappu and Sahade, 1973). Statistically, many of the apparently brightest Wolf-Rayet stars in the sky were binaries, but this result is partially biased by the increased luminosity of the system due to the bright O-type companions. Vanbeveren and Conti realized that, contrary to expectation, companions were not to be found in large numbers among fainter stars that had previously been observed only with lower dispersion. In particular, among apparently brighter LMC Wolf-Rayet stars, absorption lines due to companions were common. In the fainter LMC Wolf-Rayet stars, hardly any stars were found to have companions. Massey (1981) later investigated the true binary frequency of galactic Wolf-Rayet stars (with O companions) and found it to be ~30 percent, similar to that of O-type stars (Conti et al., 1980). This general topic is treated in more detail in Section VI.

The classification of Wolf-Rayet stars has been discussed extensively in Chapter 1 of this volume, and I will not go over all of it again here. It is basically built on the Smith (1968a) system, with an expansion to earlier and later WN and earlier WC types (van der Hucht et al., 1981). The types depend on line ratios among nitrogen ions and the appearance of He I in WN stars and on line ratios among carbon and oxygen ions and line widths in WC stars. WN subtypes range from WN2 to WN9 and WC classes range from WC4 to WC9. Vanbeveren and Conti (1980) introduced the convenient, but in hindsight possibly misleading, abbreviations of WNE for early WN's, WN2...WN6, and WNL for late subtypes, WN7 and WN8. (The WN6 group has sometimes been assigned the WNL label.) Similarly, WCE has been used for WC4...WC7, and WCL has been used for WC8 and WC9. These distinctions *may* eventually have evolutionary significance, but could also be a source of confusion in the future. For now, they should be regarded only as shorthand notations.

1. WN Stars. Smith and Kuhi (1981) have presented an atlas of optical spectra of ten WN-type stars. These photographic data were based on coudé spectrograms obtained at the Lick Observatory. The wavelength region covered was λλ 3180–7000 Å, and most subtypes were observed. Line identifications are reasonably complete, based on extending the previous work of Underhill (1968), Bappu (1973), and others.

The extensive optical spectroscopic survey of Conti et al. (1983b) lists equivalent widths for 73 galactic WN stars and 35 objects in the LMC. The strongest line is typically λ4686 He II, with line strengths up to 200 Å. Most measured lines have equivalent widths of tens of angstroms. Substantially weaker lines could not usually be measured but are present in many stars.

The line profiles were studied by Perry (1983) in his Ph.D. thesis. Line widths (at half intensity maximum) were tabulated for many lines in many stars. Although Hiltner and Schild (1966) had segregated WN subtypes into two groups—those with narrow lines, WNA, and those with broad lines, WNB—the quantitative measures of widths revealed no dichotomy. Walborn (1974) had also used suffixes, such as WN A(B), for some WN stars with intermediate widths. Although such letters are perhaps useful in designating the sharpest and broader lined WN stars, the distinction is not physical because the widths show a continuum of values.

Conti et al. (1983b) found that the emission-line strengths of WN stars typically had a dispersion of a factor of 10 among various stars, even of the same subtype. Figure 2-3, adopted

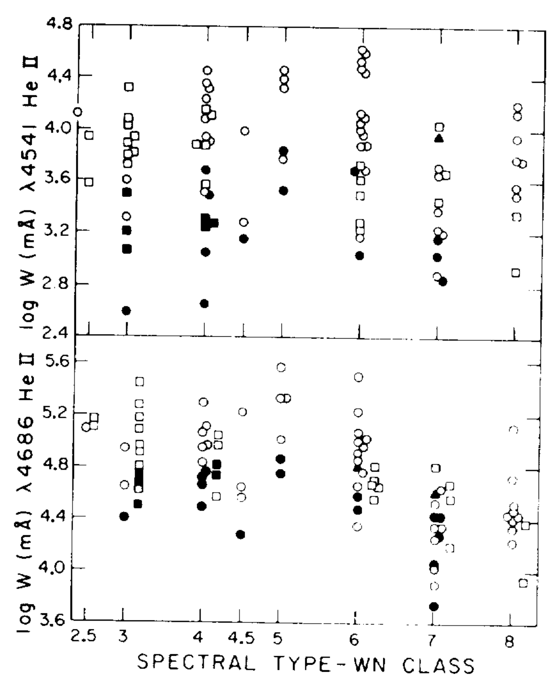


Figure 2-3. Equivalent widths of λ4200 He II for WN-type stars (from Conti et al., 1983b). Filled symbols are known binaries or stars with absorption lines: circles denote galactic stars; squares denote stars in the LMC.

from their paper, shows the equivalent widths of λ4200 He II as a function of spectral type. There was no real distinction in properties among WN stars in the Galaxy and LMC. Two effects contributed to the line-width dispersion, but they do not explain the entire range.

Those stars with O-type companions generally had the weakest emission lines, as would be expected from a continuum dominated by the companion. In addition, those WN stars that showed some evidence of hydrogen, as evidenced by the alternating stronger/weaker line strengths of the Balmer (H)/Pickering (He II) line series (see also Smith, 1973) had generally weaker emission lines. Thus, the H/He content, which must be different in some WN stars from others, even at the same spectral subtypes (Conti and Massey, 1981a; Conti, 1982a) apparently affects the wind structure in some fundamental way. Leaving aside the binaries and the WN stars with hydrogen, the remaining WN objects still evidence considerable heterogeneity in their spectra. As will be shown later, this is also true of their masses (Section VI) and luminosities (Chapter 3), but a relationship among these parameters has not yet been established.

Perry (1983) found that most emission lines (80 percent) in most stars (80 percent) could be closely fitted with gaussian profiles. Some other lines were parabolic, and the remainder were similar to profiles that would be composed of two separate gaussians—one a stronger and narrower profile and one weaker but broader. With our still-limited knowledge of the physical structure of the winds of WN stars, Perry was not able to conclusively constrain the parameters to derive meaningful results, but could assert that there were various reasonable distributions of ionization, density, and velocity with height that could produce gaussian profiles in optically thin cases. Optically thick lines would readily lead to parabolic profiles. Two-gaussian forms come about naturally in cases in which the velocity law is not monotonic, although other explanations remain possible.

Conti et al. (1983b) also found that quantitative measurements of the N III, N IV, and N V lines, which are the major determinants of the subtype, indicate a real change in the ionization state of the stellar wind from one subtype to another. However, some overlap occurred in ratios between some adjacent subtypes. Another way of expressing this is to say that there are too many subclasses, considering the quantitative change in N III/N IV and N IV/N V ratios. In WN6–WN8 stars, the appearance of He I also plays a role in the classification. Some ambiguity exists in the WN subclass for these WNL stars. On the average, WN7 stars seem to have weaker emission lines than the adjacent subtypes. Some of them may have different intrinsic properties, particularly if they tend toward the "transition" appearing case. One-third of all WN stars contain hydrogen from the Balmer/Pickering decrement. The other two-thirds have no evidence of hydrogen, but whether the H/He ratio is 0.1 or 0 cannot be determined with the available data.

C IV λλ 5801,5812 lines are visible in quite a number of WN stars. C III λ4650 (and λ5696) lines are not seen; those few objects in which such lines are present are labeled WN + WC (van der Hucht et al., 1981; Conti, 1982b) and might represent the spectra of two stars (e.g., the case for R140 of 30 Doradus in the LMC (Phillips, 1982)) or may be objects in "transition" from WN to WC composition. Conti et al. (1983b) measured the C IV features in many WN stars and discussed these issues. They find that, in most WN stars, the C IV/N IV ratio indicates nitrogen enhancement at the expense of carbon (Chapter 3, Section VII).

Vreux et al. (1983) have presented a catalog of near-IR spectra of southern galactic Wolf-Rayet stars. These data supersede previous scattered results in the literature because they have higher resolution and signal-to-noise ratios. They identify lines from common ions of helium, carbon, and nitrogen in 12 WN stars. Equivalent widths up to a few hundred angstroms are found for λ10124 He II. Most lines are a few tens of angstroms, and weaker lines could not readily be measured.

Hillier et al. (1983) determined emission-line profiles in the far-IR region for the WN5 star

HD 50896. These are the first solid quantitative spectroscopic data for a WN star in this spectral range. In his thesis, Hillier (1983) makes a detailed analysis of this star, modeling these lines. Low-resolution spectroscopy of HD 50896 from 1.4 to 2.5 μm was also published by Williams et al. (1980).

Oxygen lines are not usually readily seen in WN-type stars, but HD 104994 is an exception (van der Hucht et al., 1981, and references therein). The λλ 3811,3834 O VI lines are observed, and other transitions for this ion are found in the near-IR (Vreux et al., 1983).

Before the launch of IUE, ultraviolet observations of Wolf-Rayet stars were generally restricted to low resolution or a small sample of the brightest objects. The OAO-2 satellite was used to obtain the first UV spectrum of a WN star, HD 50896 (Smith, 1973). Willis and Wilson (1978) reported 35 Å resolution measurements of three WN stars with the TD-1 satellite, and six Wolf-Rayet stars were observed with Skylab (Henize et al., 1975). Johnson (1978) gives an atlas of Copernicus spectra for two WN stars.

The IUE satellite has revolutionized the study of Wolf-Rayet stars. An atlas of high-resolution data (0.1 to 0.2 Å) for nine "single" WN stars has been presented by Willis et al. (1986). Spectral tracings and line identifications are found in their paper. The following species of ions are observed in these nine WN stars: He II, N II to N V, C III to C IV, and O III to V, and ions of Mg, Al, S, P, Fe, and Zn. The dominant lines are, of course, those of He II and nitrogen (λλ 1548,1550 C IV is strong).

Preliminary analyses of these data, insofar as they pertain to the stellar wind, have been presented by Willis (1981, 1982). Considerable low-resolution (6 Å) IUE data have been acquired for galactic and LMC WN stars. A few results have been reported (e.g., Garmany and Conti, 1982; Nussbaumer et al., 1982).

2. WC Stars. There is no equivalent published Smith and Kuhi atlas for WC stars, although similar spectroscopic data exist (Smith and Kuhi, private communication). An extensive optical spectroscopic survey of WC stars was undertaken by Conti which is similar to that reported for WN stars. A. Torres studied these data for her Ph.D. thesis, and I will present some published and preliminary results here.

Torres and Conti (1984) studied the 12 known members of the WC9 subclass. These stars have the narrowest emission lines and the most complicated spectra, being dominated by various C III transitions, with C IV and C II also present. They found a surprising similarity in line strengths and widths among this type. The homogeneity here is to be contrasted with the heterogeneity among the WN stars.

Some preliminary results from Torres (1985) include finding good correlation, as expected, between the WC subtypes as defined among the C III, C IV, and O V line ratios. The earlier types have broader lines. However, 10 percent of the WC stars in the Galaxy do not fit this relation because they have somewhat broader lines. Torres et al. (1986) noted that two of the WC stars having very broad lines for their subtypes, WR 125 and HD 193793, are also listed as nonthermal radio emitters and anomalously strong radio sources. The cause of these features is not known, but the exceptionally broad lines may be a key to our understanding of phenomena at other wavelength regions. Nearly all WC stars in the LMC have line widths more extreme than WC stars in the Galaxy. As has been known for some time, the LMC WC stars are all early-type stars. In fact, they show extremely high ionization states, and according to Torres et al., are more properly called WC4. Our understanding of the winds in WC stars is not yet clear enough to give us a satisfactory reason for the dichotomy in the WC stars in the LMC as compared to the Galaxy, nor for the absence of the later WC types.

The classification line ratios among WC subtypes differ sufficiently because not much overlap occurs between adjacent classes. The line strengths scatter among WC stars of the same subtype, although not as much as for WN stars. Figure 2-4 shows this for the λ5806 C IV blend, adopted from Torres et al. (1986). This figure was plotted with the revised WC types

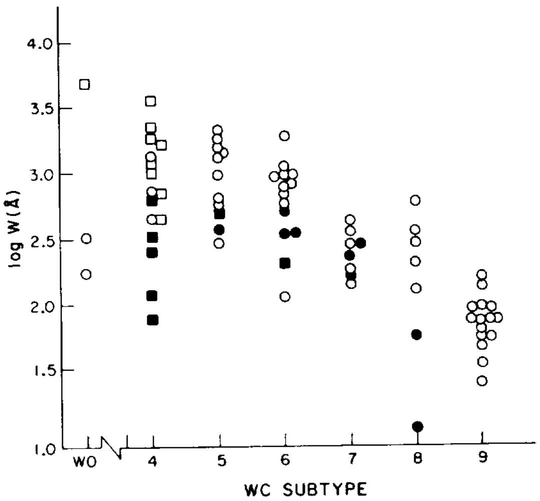


Figure 2-4. Equivalent widths of λ5812 C IV for WC-type stars (from Torres et al., 1986). Filled symbols are known binaries or stars with absorption lines: circles denote galactic stars; squares denote stars in the LMC.

from that paper. The stars with companions have weaker emission, as is expected from the contamination in the continuum.

The strongest features in WC stars are the λλ 4650/4686 C III/He II blend, with equivalent widths to 1000 Å being measured. The C IV doublet at λλ 5801,5812 is also quite prominent, ranging to 400 Å in equivalent width. There is no evidence of hydrogen in any WC star (Nugis, 1975), although blending of carbon ions with the Balmer series makes that conclusion less certain for WC stars than for WN objects. No lines of nitrogen are seen in WC stars. In particular, the λλ 4634,4640 N III, λ4057 N IV, and λλ 4603,4618 N V lines, which are prominent in WN stars, are not visible.

Vreux et al. (1983) has given near-IR spectrometry for 14 WC-type stars. Williams et al. (1980) published lower resolution data for three WC stars, extending to 2.5 μm. The strongest feature is due to transitions in λλ 9696–9718 C III, where equivalent widths to 1000 Å were found. Most lines have tens of angstroms equivalent widths, and weaker features are present but could not be measured. These lines are mainly due to He II, C II, C III, and C IV. A few oxygen lines are also seen. No nitrogen lines are found in the IR spectra of WC stars. In particular, the λλ 7103,7128 N IV line, which is very prominent in WN stars, is not present. A strong 3.3-μm emission line is found in a few stars and was shown by Williams et al. to be due to C IV transitions.

Pre-IUE FUV spectroscopic observations of WC stars were scarce in number. Using the OAO-2 satellite, West (1972) observed γ^2 Vel, WC8 + O9I. Willis and Wilson (1978) reported TD-1 satellite data on six WC stars, and Henize et al. (1975) published data concerning six WC stars from the Skylab mission. γ^2 Vel was bright enough to be the subject of several rocket-launched experiments: Carruthers (1968) showed the first low-resolution (2 Å) results, and Burton et al. (1975) and Willis and Wilson (1976) gave high-resolution (0.3 Å) data from Skylab experiments. Johnson (1978) has published an extensive FUV spectral atlas at 0.2 to 0.4 Å resolution that gives data from λ946 to λ3182 Å for γ^2 Vel.

The high-resolution IUE atlas of five "single" WC stars of various subtypes (Willis et al., 1986) gives the most complete line lists and identifications in the wavelength range λλ 1150 to λλ 3250 Å. The following ions were found: He II, C II–C V, O III–O V, Mg IV, Al II–IV, Si II–IV, P II–V, S III–IV, and Fe III–VI. Although nitrogen lines were searched for, none were found. In particular, the resonant λλ 1238,1240 N V features, so prominent in WN stars and very strong in all O-type stars, were not found. The λλ 1486,1718 N IV lines, also very strong in WN stars, were not present. (Features near the latter wavelength are due to silicon ions (Willis, 1981).) The absence of nitrogen in WC stars, over all observed wavelength regions, is noteworthy and will be considered again later (Chapter 3, Section VII).

3. WO Stars. The WO class has recently been introduced by Barlow and Hummer (1982) to designate those few apparently Population I stars in which the dominant feature in the optical regions are the λλ 3811,3834 O VI lines. These lines are seen in many WC stars (also in HD 104994, a WN3 star), but their strength in

WO stars is unprecedented and comparable to the He II and C IV lines. Strong O VI lines are often also found among central stars of planetary nebulae (Smith and Aller, 1969). Sanduleak (1971) listed five O VI stars not associated with planetary nebulae, including one in the SMC, Sand 1 = Sk188, and one in the LMC, Sand 2 = FD73 = BR93. These latter stars are sufficiently luminous to be associated with a parent massive star population.

Barlow et al. (1980) have discussed optical observations of Sand 3 = WR72, the star with the narrowest lines. They identify lines from O VII, O VIII, and C V, implying excitation energies in excess of 800 eV. Oxygen seems to be more prominent than carbon, and He II is present but relatively weak. Barlow et al. (1980), Barlow and Hummer (1982), and Sanduleak (1971) suggested that WO stars might be more highly evolved than WC stars, given the apparent abundance anomalies.

Through Infrared Astronomical Satellite (IRAS) observations, van der Hucht et al. (1985) have discovered a dust shell around Sand 3. They conclude that this star is the central star of a planetary and not a Population I object. (Note that Sand 3 is far out of the galactic plane: b = +14°.) Barlow and Hummer (1982) had also suggested this possibility. The IRAS did not detect dust around the other WO stars, but the limits are not particularly interesting. Certainly the two Magellanic Cloud WO stars are Population I objects; the two other galactic ones may be, but more data are needed. All four remaining Sanduleak WO stars have much broader lines than Sand 3, a common characteristic of Population I *vis a vis* Population II Wolf-Rayet stars (Smith and Aller, 1971).

Thevenin and Pitault (1982) have discussed a 2 Å resolution yellow-red spectrum of Sand 5. They identify an O VI line at λ5291. C III λ5696 is not present, although lines of C IV and He II are; thus, the excitation class is very high. Thevenin and Pitault point out the extreme line width of Sand 5 and its similarity to Sand 4, while also noting that Sand 3 is more properly related to Population II Wolf-Rayet objects.

4. Transition WN Stars. Walborn (1974) illustrated a number of late-type WN spectra, along with an Of star, HD 93129A. He pointed out the morphological similarity among these Of and WN spectra. Conti (1976) proposed an evolutionary sequence connecting these data, suggesting that Of stars evolved to late-type WN stars with mass loss. The several late WN stars that contained hydrogen, some (e.g., HD 92740, HD 93131, and HD 93162) even showing the upper Balmer lines in absorption (Niemela, 1974; Conti et al., 1979), would then be transition objects between the Of and classic WN types. I will return to the evolutionary picture later in Chapter 5.

For now, let us examine merely the transitional spectroscopic nature of some luminous Of/WN stars. An Of star has primarily an absorption-line spectrum with relatively narrow emission lines of He II λ4686. A WN star has primarily a broad emission-line spectrum, with absorption lines (in the Balmer series) in a very few cases. What then of stars with an absorption-line spectrum and a broad λ4686 He II? Two recently identified examples are Sk−71°34 and Sk−67°22 (Conti and Garmany, 1983; Walborn, 1982a) in the LMC. These could be two stars (each), in which case the types are WN3 + O4 and WN4.5 + O3, respectively. As yet, no radial-velocity studies of these objects have been done to confirm or deny duality. On the other hand, they could also be single stars, in which case, according to the listed authors, they represent mixed O4f/WN3 and O3If/WN6 subtypes, respectively.

Other luminous objects in and near 30 Dor with transitional spectra are listed by Conti (1982b) and Walborn (1984b). HDE 313846, called by Hutchings (1979) the "most extreme Of star," was classified WN9 by van der Hucht et al. (1981). Walborn (1977, 1982a) has discussed other stars in the Galaxy and LMC which show transitional spectroscopic properties. Breysacher (1981) has also considered the classification of several of these stars. Some galactic objects have been shown to be single

(Conti et al., 1979) through radial-velocity studies. HD 15570 has transitional properties in its UV spectrum (Willis and Strickland, 1980), although its optical classification is unambiguously O4f. This star is apparently single.

Luminous objects with spectroscopic properties of both Of and WN stars thus exist, and at least some are single. Stahl et al. (1983) have reported the remarkable behavior of the LMC supergiant, R127 (= HDE 269858), which had been classified as an Of star by Walborn (1977, 1982a): O I af pe extr or, alternatively WN9–10) but had suddenly brightened by ~0.75 magnitude. This was accompanied by a dramatic change in the spectrum to a later spectral type of a B emission-line supergiant. The N III and He II emission lines characteristic of the Of phase had disappeared, to be replaced by one in which the wind shows a lesser velocity but denser outflow (Walborn, 1948a). Details are contained in the Stahl et al. paper, in which they suggest that the bolometric magnitude has not changed and the brightening is due to an increase of the effective stellar radius. Will other transitional Of/WN stars show similar behavior? Only time will tell, but this luminous class clearly bears extensive and continuous monitoring.

Walborn (1982a) pointed out that the narrow emission-line spectrum representing nebular material surrounding the transition star, Sk–67°266, has a clear enhancement of nitrogen. Perhaps this material was ejected in a relatively recent outburst, analogous to that of R127. Nitrogen enhancement would be expected in CNO equilibrium material mixed up from the nuclear reactions in the stellar interior. (See, for example, the case of η Carina by Davidson et al., 1982.)

IV. NUMBERS AND DISTRIBUTION

A. O-Type Stars

1. Galactic Volume-Limited Sample. These very luminous O-type stars (M_v ~ −3.5 to −7.0) can be seen to great distances. Their positions with respect to the Sun can be found from their stellar coordinates and their distances. The latter quantity is estimated by the method of spectroscopic parallaxes; a stellar spectrum is taken, and the use of an M_v-spectral-type relation enables one to estimate the absolute magnitude from it. Photometry, combined with a knowledge of the intrinsic color (see Part I of this volume, but note that $(B-V)_0$ for O stars are similar) and the adoption of a normal reddening law (e.g., Savage and Mathis, 1979), leads in a straightforward manner to an estimate of the distance. The greatest uncertainty in this case is the M_v-spectral-type relation. This will be discussed in more detail in Chapter 3 (Table 3-1), but for now, note that the accuracy is $\pm 0\overset{m}{.}5$ for a single star.

The publication of the *O-star Catalogue* of Cruz-Gonzalez et al. (1974) greatly aided discussions of distributions of O-type stars. These authors collated data from the literature to list 664 O-type stars known in the Galaxy at that time. Lequeux (1979) used these data in one of the first investigations of the O-star statistics. Garmany et al. (1982) counted the O stars of the *Catalogue* in concentric rings and found that the number counts went roughly as a power 2 law (as they should for a flattened disk population) out to roughly 2.5 kpc from the Sun, where the values began to turn over. Thus, the distribution was roughly complete to this distance, and the concept of discussing a volume-limited sample for O stars began. Bertelli and Chiosi (1982) also discussed the O-star distribution near the Sun.

Garmany et al. (1982) included additional spectral types of some *Catalogue* O stars with photometry but without known spectral types, and a number of additional O stars have since been added to the *Catalogue* list. A machine-readable list is being kept current by Garmany in Boulder, Colorado. The initial stellar evolution models for massive stars (see Chapter 5, Section II) indicated that core hydrogen-burning ceased roughly at the O-star/B-star boundary, hence Garmany et al. (1982) could argue that only the O-star statistics were needed to estimate the initial numbers of massive stars. However, it became increasingly clear from statistical arguments that the upper main sequence is "wider" than models indicate, as argued by Stothers and Chin (1977), Chiosi et al. (1978), Cloutman and Whitaker (1980), and Bressan et al. (1981). Meylan and Maeder (1982) also pointed out an observational confirmation of this in an analysis of cluster magnitude diagrams. The "main sequence" of core hydrogen-burning thus extends to the later B supergiant regime for massive stars (Conti et al., 1983a; Conti, 1984). A current catalog of known B supergiants (adopted from Humphreys and McElroy, 1984) is also kept current.

Figure 2-5 is a plot of the distribution of O stars and luminous B supergiants, projected onto the plane of the Galaxy (Sun at the center). This updates the analogous figure in Garmany et al. (1982). (See also Van Buren, 1985.) The 795 stars in this sample (some are overplotted in the figure), based on initial masses $\gtrsim 20$ $M_\odot$, are shown to 3 kpc, but the population is possibly not completely identified in the outermost annulus. It is instructive to compare this figure with its historical antecedent of tens

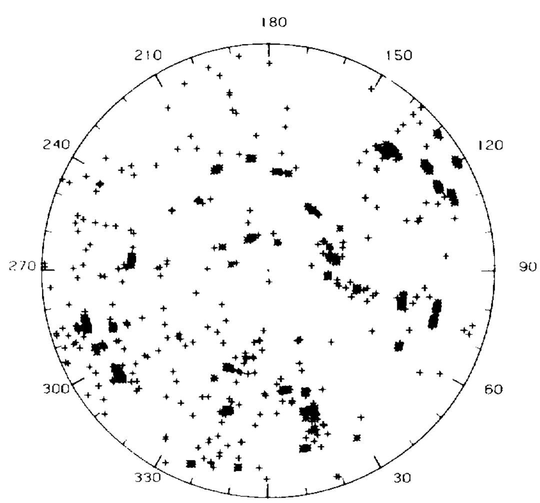


Figure 2-5. Galactic longitude distribution of O stars (plus signs) and B supergiants (asterisks) projected on the plane of the Galaxy. Only stars within 3 kpc and with inferred masses larger than 20 $M_\odot$ are shown.

of OB stars (e.g., Morgan et al., 1950). Three spiral-arm features are seen: (1) toward the galactic center and around to longitude 300° is the Scorpio-Sagittarius arm; (2) in the 85° direction is the Cygnus arm, in which the Sun appears to be presently located, which can be followed also toward 270 to 280° in the Carina direction; and (3) outward from the Sun, particularly between 110 and 140° longitude is the Perseus arm, which is not readily followed toward 210°. I must say that, if we didn't know our Galaxy was a spiral, this diagram would be of no use to convince the skeptic. But the "arms" described here are well known to radio astronomers who study 21-cm radiation. The OB stars are coincident with the H I peak emissions, but are greatly spread out in dimension from them. I do not believe that this represents inaccuracies in the distance estimates and "spectroscopic parallaxes," but rather that OB stars in our local vicinity have a rather broad physical extent compared to the neutral hydrogen.

Stepping back from the difficult perception of spiral structure in Figure 2-5, it can also be seen that there is a negative gradient in the number density of OB stars with galactocentric distance; somewhat more are seen toward the galactic center. (Arbitrarily dividing the volume in half from ℓ = 90 to 270° yields a number ratio inward/outward of 511/284 = 1.80.) This gradient is qualitatively in accord with galactic mass models in which more matter is found toward the center.

Updating the results of Conti et al. (1983b), we divide the O-star population into stars more and less massive than 40 $M_\odot$, using the evolution tracks of Maeder (Chapter 5, Section II). Figure 2-6 shows the OB stars with initial masses of 20 to 40 $M_\odot$, from the previous listing. The 550 stars in this diagram have a similar morphological appearance to Figure 2-5 (i.e., they are patterned the same way and show a small gradient in the numbers with galactocentric distance). The number ratio inward/outward is 334/216 = 1.65. On the other hand, Figure 2-7 shows the remaining 245 OB stars with initial masses $\gtrsim 40$ $M_\odot$. Here, the pattern

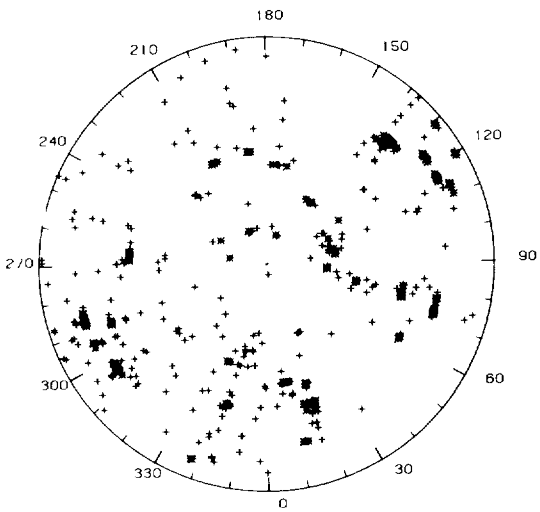


Figure 2-6. Galactic longitude distribution of O stars and B supergiants with inferred masses $M < 40\ M_\odot$. Symbols and criteria as for Figure 2-5.

is distinctly different from Figures 2-5 and 2-6. A substantial gradient can be seen in the distribution; relatively few massive OB stars are found in the galactic anticenter region. The number ratio inward/outward is 177/68 = 2.60. These figures show an important result; there are relatively more massive stars ($\gtrsim 40\ M_\odot$) toward the galactic center than toward the anticenter region, compared to less massive types ($20\ M_\odot, \lesssim M \lesssim 40\ M_\odot$). This result depends only on stellar coordinates and is uncontrovertible, unless we are missing the detection of massive and *very luminous* stars toward the anticenter or *large numbers* of less luminous stars toward the center. Both of these arguments seem unlikely. It is tempting to associate this massive star gradient with galactocentric distance, but it would be prudent to merely restate our conclusion as follows: relatively *more massive stars are found in the inner Scorpio-Sagittarius arm than in the outer Perseus arm.*

Parenthetically, it might also be noted that the "spiral arms" for the massive stars in Figure 2-7 are broad with respect to the interarm spacing. This suggests to me that the spiral structure we perceive in other galaxies is predominantly determined by the optical appearances in the H II regions, which are excited by groups (associations) of massive O-type stars. Detailed stellar studies of nearby spiral galaxies would help to address this point because it is difficult to get a "world view" of our own system. Nearly all the OB stars in Figure 2-7 are indeed in associations. Section IV.B will show that the galactic distribution of Wolf-Rayet objects is similar to the massive stars of Figure 2-7, thus showing an intimate evolutionary connection (Chapter 5, Section II).

2. z-Distribution of O-Type Stars. We can find the z-distribution of stars in our Galaxy from their distances and galactic coordinates. Figure 2-8 shows the z-distribution for the stars of Figure 2-5, for which the distances are within 3 kpc. This figure indicates an approximate scale height of ~50 pc from the "thickness" of the distribution in z. Note that the scale height does not appear to be uniform with galactic longitude, but somewhat larger in the fourth quadrant ($\ell \sim 270$ to 360°). Also, the centroid of the O-star distribution does not follow the galactic equator, but can be above or below it by 50 pc (e.g., +50 pc at $\ell \sim 340°$

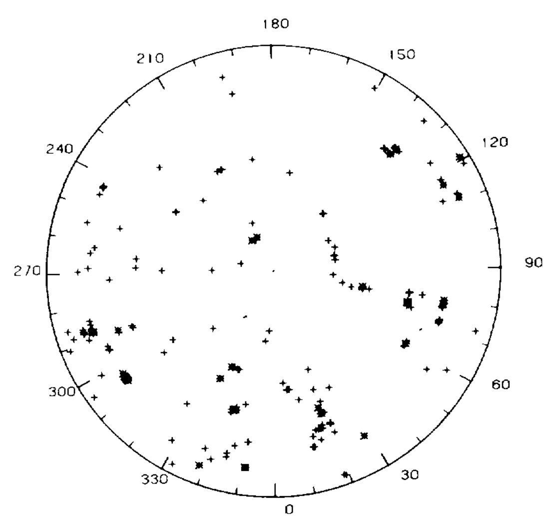


Figure 2-7. Galactic longitude distribution for O stars and B supergiants with inferred masses $M < 40\ M_\odot$. Symbols and criteria as for Figure 2.5.

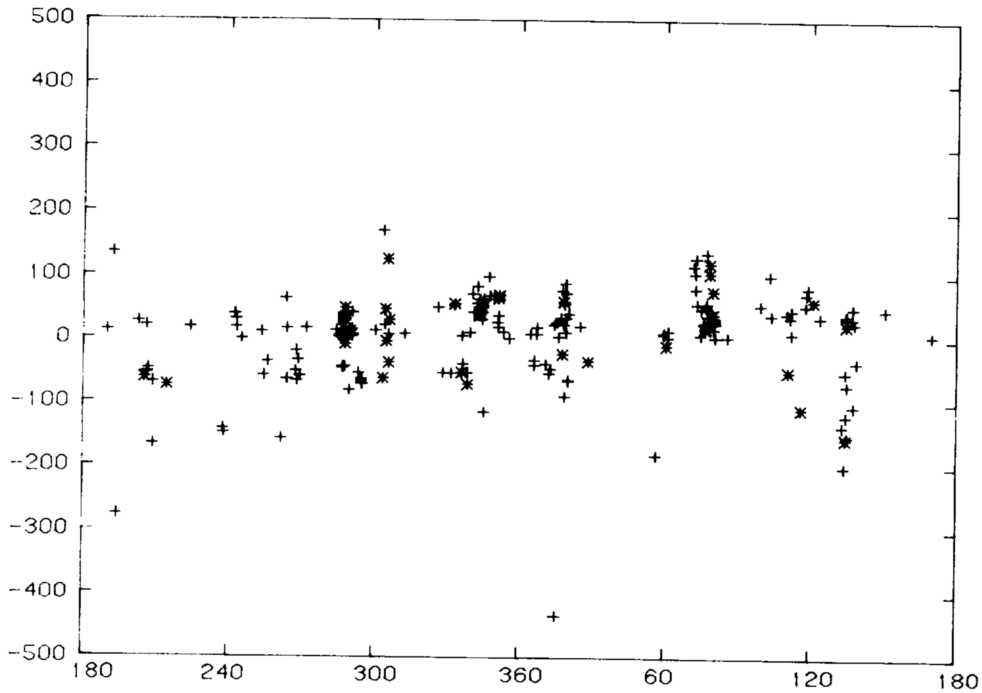


Figure 2-8. Galactic latitude distribution (in terms of z-distance) for stars of Figure 2-6, using same symbols and criteria.

in Scorpio and -50 pc at $\ell \sim 200°$ in the anticenter). A few nearby O stars are found far from the warped plane and are in anomalous positions. Some are high-velocity objects (Section V).

3. O-Stars in the Magellanic Clouds. Objective prism surveys of the Magellanic Clouds have revealed numerous hot luminous OB-type stars. Sanduleak (1969) has provided a list, with finding charts, of 1272 OB-type stars in the LMC; a similar survey of stars in the SMC is given by Sanduleak (1968). Henize (1956) has listed Hα emission-line stars in the Magellanic Clouds, nearly all of which are early types. Bohannan and Epps (1974) have made a deeper Hα survey of the LMC; a similar deep survey of the SMC is not yet available. Important catalogs of LMC members are those of Brunet et al. (1975), Fehrenbach et al. (1976), and Rousseau et al. (1977), which contain 1791 LMC members. An analogous SMC catalog is that of Azzopardi and Vigneau (1975, 1982), which contains 524 members. These catalogs list stars to $M_v \sim 14^m$, but are probably not complete to quite this magnitude. With a modulus of ~18.6 for the LMC and ~19.1 for the SMC, these catalogs will not reach all the late-type main-sequence O stars.

Although most of the stars in these catalogs have photometry, only a small fraction have slit spectra and have been classified as to MK type. Of the stars in the Rousseau et al. catalog, 266 have MK types, for the most part B-type supergiants. Crampton (1979) has subsequently obtained classification spectra of 20 additional O stars in the LMC. Shore and Sanduleak (1984) have discussed slit spectra (and IUE observations) of 19 early-type B emission-line objects in the LMC and two in the SMC. The Azzopardi and Vigneau (1982) SMC catalog lists 253 stars with MK types, again mostly B supergiants, and 19 O types.

Conti, Garmany, and Massey are engaged in a major project to identify and classify additional O-type stars in the Magellanic Clouds.

As has been emphasized by Massey (1985), the *UBV* colors of the earliest type stars are not very different. The difference in $(U-B)_0$ between O5 and B0 main-sequence stars is only $0\overset{m}{.}09$ (FitzGerald, 1970); photometric uncertainties, and questions of the reddening laws in the Magellanic Clouds, mean that one cannot distinguish among O stars by optical photometry alone. One must have spectra: B0 main-sequence stars have masses $\approx 15\ M_\odot$, and early O-type stars have masses ranging to 100 $M_\odot$. Thus, classification of hot stars is critical. Fortunately, B0.5 and later-type main-sequence stars and supergiants begin to have appreciably different intrinsic colors and can be roughly classified on the basis of reasonably accurate photometry alone. Membership of these stars in the Magellanic Clouds also provides luminosity estimates.

Do the Magellanic Cloud catalogs contain all the O-star candidates? Probably not, because they do not go faint enough. Also, many hot stars are found in compact groupings or associations (e.g., Walborn, 1973a) which, in many cases, are too crowded to have enabled objective prism spectra to have been obtained. Lucke (1972) and Lucke and Hodge (1970) have identified 122 associations in the LMC (the L-H associations), of which only the brighter stars are listed in the LMC catalogs. A similar identification in the SMC has recently been given by Hodge (1985). Lucke has also provided photographic photometry for the brighter stars (to $M_v \approx 16^m$) in the L-H associations. Approximately 2000 "blue" stars are listed in his thesis, but even so, other objects could not be measured because of associated nebulosity. Of the 2000 stars, many will be B supergiants, but possibly ~ 1000 O stars remain to be found in the L-H associations in the LMC. According to Hodge, there are approximately 70 stellar associations in the SMC (NGC 346 being the most obvious example), but their stellar membership is sparse. Perhaps 50 O stars could be in these groupings but not incorporated in the Azzopardi and Vigneau (1982) catalog.

Our ongoing spectroscopic identification program has revealed a total of 134 O stars in the LMC and 63 in the SMC at this writing. The totals are still incomplete, and our work is continuing. It is therefore still premature to discuss the numbers and distributions of O stars in the Magellanic Clouds, even though these data are well in hand for Wolf-Rayet stars, as we shall see in Section IV.B. Our best guess at the present time is that there are 500 to 2000 O stars in the LMC and 50 to 200 in the SMC. No discussion of the Magellanic Clouds would be complete without mentioning the Hodge and Wright (1967, 1977) atlases of these nearby galaxies, which are absolutely essential to investigations of stellar distributions.

4. O Stars in Other Galaxies. We know that O stars are prevalent in the spiral-arm populations in other galaxies because they are responsible for producing the H II regions which are such a dominant feature of late-type galaxies and irregulars. Among the local group galaxies (M31, M33, NGC 6822, and IC 1613), with distance moduli typically $\sim 24^m$ to 25^m, the brightest normal O stars will appear at $m_v \sim 18^m$ to 19^m. Individual spectroscopy is difficult, but not impossible. Broad surveys with slit spectra are clearly beyond the capability of ground-based equipment, but UV imaging telescopes would make specific identification a tractable problem.

A handful of O- and Of-type stars have been identified in M33 by Conti and Massey (1981b) and Massey and Hutchings (1983) in their investigations of exciting objects of the giant H II regions. Luminous blue stars, nearly all B- and A-type supergiants, have been specifically identified in Baade's field IV in M31 (Humphreys, 1979), in NGC 6822 and IC 1613 (Humphreys, 1980a), and in M101 and NGC 2403 (Humphreys, 1980b). Humphreys and Sandage (1980) have also discussed the stellar content of M31: all the brightest blue stars in these galaxies are B and A supergiants, as expected, since the massive O stars have substantial bolometric corrections and are therefore fainter in M_v.

5. Indirect Inferences: H II Regions. H II regions are produced by Lyman continuum

(LyC) photons emitted from hot stars, predominantly O types. A large amount of literature exists. Investigators such as Güsten and Mezger (1983) and Israel (1980) have used the radio emission measurements from these H II regions to discuss the number of O-type stars present. This measured radio emission is essentially related to the number of recombinations of electrons with ions; the recombination rate is equal to the LyC ionization rate under the reasonable assumption that the H II region is in equilibrium with its central exciting star(s). This is probably satisfactory except in the very earliest stages when the star(s) first "turn on."

The LyC photon number inferred from the foregoing relationship must be corrected for absorption by dust in the nebulae or for the fraction which escapes without ionizing hydrogen. With these corrections, which act in opposite directions, one can then estimate the number of LyC photons emitted by the stars themselves. At this point, one must appeal to stellar atmospheric models to relate the LyC photons present to the stellar effective temperatures and radii. The atmospheric models used have generally been based on the Mihalas (1972) non-LTE code, which is plane parallel and does not include the effects of the winds or the possible role of backscattering. Panagia (1973) tabulated the relationship between emitted LyC photons and stellar spectral types. Stasinska (1982) has tabulated a series of nebular models which are commonly used to predict the structure of the gas. Borsenberger and Stasinska (1982) made a crude attempt to account for situations in which the exciting stars had a helium or metal content different from solar and, therefore, a different internal structure. With the stellar models used, the predicted effects are minor, but they did caution that, in real stars where line-blanketing needs to be considered, the effects might be substantial. These models are the basis of nearly all the work in which LyC measurements are used to predict how many O stars are present.

An O-star spectral-type/mass relation, which is based on stellar structure models, must be used in addition to the foregoing steps as a final procedure to obtain information on the initial mass function (IMF), or star formation rate (SFR), of massive stars. For example, Güsten and Mezger (1983) thoroughly discuss all the detailed procedures in order to derive the star formation rate in the Galaxy, based on measures of giant H II regions and extended low-density H II regions. Israel (1980) uses radio observations of H II regions in external galaxies to describe the number of O stars present. Viallefond (1984a, 1984b) discusses the IMF and SFR in the ionized nebulae of M33. Other examples are quoted in these references and elsewhere.

The foregoing procedures should be conceptually correct because the O stars certainly produced the dominant LyC luminosity. However, the steps in estimating the T_{eff}, thus attempting to count how many O stars are present, or what masses are inferred, are very indirect; corrections are needed in the data (the fraction of photons observed to those produced) and in the models (atmospheres and structures). The quoted results are interesting; for example, Israel (1980) predicts that 3×10^3 exciting O stars are present in the SMC, and 1.3×10^4 are present in the LMC. Although some of these implied objects may be B0–B1 main-sequence stars that have not been identified in our star counts, in both cases, these predictions appear to be factors of 4 to 10 too high. It is well known that the nebular excitation parameter (i.e., O III/Hβ ratio) goes up with galactocentric distance in spiral galaxies (Searle, 1971), and the properties seem to indicate higher excitation farther from the galactic center. This has come to be inferred as indicating that the temperature of the exciting stars increases likewise, as does their mass. For example, in M33, Viallefond (1984a, 1984b) suggests that the number of massive stars should increase with galactocentric distance in M33, the reverse of what star counts in the solar vicinity would imply. My view is that his conclusion may be an artifact of the inadequacy of the models.

In particular, we now know that stellar winds are important features of O stars because their structure and strength affect the underlying photospheric structure (Hummer, 1982; Abbott and Hummer, 1985). We will see in Chapter 3 that the current atmospheric models give roughly the correct *LyC* flux—the derived T_{eff} from the so-called "Zanstra" method agree reasonably well with other methods for "normal" composition exciting stars in the solar vicinity. However, it seems to me entirely possible that, if one has stars with metal content substantially different from our vicinity (e.g., the SMC and other regions of other galaxies), their wind-blanketing and emergent *LyC* photon number could be considerably different from normal abundance stars of the same T_{eff}. One could readily imagine that, as the metal abundance goes down, the blanketing also decreases, and more *LyC* photons are found at a given T_{eff}. Thus, one would infer that there are more exciting stars than are actually present in the SMC and conclude that there was a gradient of hotter stars with increasing galactocentric distances, rather than the reverse as shown by the number counts. All this must remain speculative until better atmospheric models are available. For now, prudence must remain the best course of action on conclusions from indirect inferences of numbers of O stars.

B. Wolf-Rayet Stars

1. Galactic Volume-Limited Sample. In our Galaxy, 160 Wolf-Rayet stars are known—159 from the Wolf-Rayet *Catalogue* of van der Hucht et al. (1981), plus a new WC9 discovered by Danks et al. (1983). The van der Hucht et al. *Catalogue* also gives complete references concerning the discovery of these stars and individual work through mid-1981. Spectrophotometry of some fainter stars has since been added by Massey and Conti (1983a) and Lundström and Stenholm (1983, 1984a).

Hidayat et al. (1982), using the Wolf-Rayet *Catalogue*, discussed the galactic longitude and latitude distributions of these stars. They pointed out the peculiar distribution of subtypes with galactocentric distance, relatively more WC stars being found toward the galactic center, particularly WC9 subtypes. Their distances of Wolf-Rayet stars were determined from photometry and a spectral-type M_v relationship given in the *Catalogue*. This will be discussed in more detail in Chapter 3, Section I.

Roberts (1962) drew attention to a peculiar galactic longitude distribution of Wolf-Rayet stars—a "zone of avoidance" in the galactic anticenter region. The same anomaly, which can be better described as a steep gradient in Wolf-Rayet numbers with galactocentric distance, is found among the massive O-type stars (Garmany et al., 1982; Conti et al., 1983a).

Updating the latter paper using the spectral-type M_v relation in Chapter 3, I show in Figure 2-9 the distribution of Wolf-Rayet stars in our Galaxy to 3 kpc. This diagram is probably essentially complete to this distance, although individual stars near the boundary could move out, or others in, given the statistical uncertainties in the M_v calibration. The overall similarity of the distribution of Wolf-Rayet stars to the massive OB stars (Figure 2-7) is striking. Relatively more Wolf-Rayet stars are found in the inner Scorpio-Sagittarius arm, and hardly any in the outer Perseus region, analogous to the massive OB-star distribution. It seems very clear that Wolf-Rayet stars and massive OB

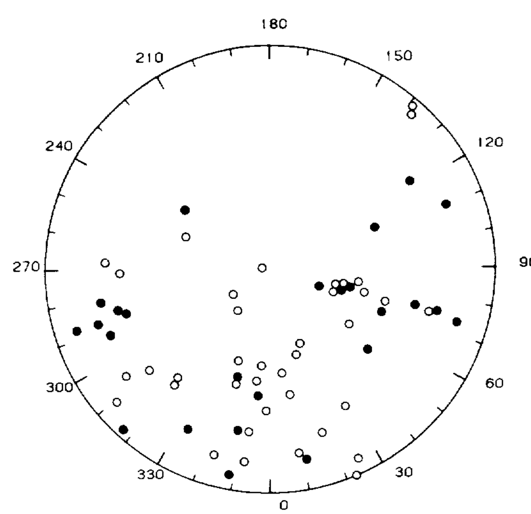


Figure 2-9. Galactic longitude distribution for Wolf-Rayet stars within 3 kpc: filled circles denote WC stars; open circles denote WN stars.

stars are intimately related. The dividing line at 40 $M_\odot$ (Conti et al., 1983) is uncertain by perhaps $\pm 5\ M_\odot$ (probable error). It would be extremely interesting to follow such distribution arguments into other galaxies of the local group, but as we have seen, the OB-star populations there are not yet known.

Although the Wolf-Rayet stars to 3 kpc are probably complete, the OB-star population is only known well to 2.5 kpc. Inside the larger volume are 245 massive OB stars and 63 Wolf-Rayet objects, giving an OB/WR ratio of 3.9. We will make use of this value later when discussing evolutionary scenarios. The ratio of WN/WC types in the solar vicinity is roughly 1 over our 3 kpc volume-limited sample. According to Figure 2-9, there are relatively more WC than WN types toward the galactic center. Unfortunately, a "complete" sample cannot be trusted to larger distances. There is a hint that the WN/WC ratio increases with galactocentric distance, although we are dealing with small-number statistics. All WN and WC subtypes are found in the solar vicinity.

Lundström and Stenholm (1984b) and Humphreys et al. (1985) have discussed the membership of Wolf-Rayet stars in galactic associations. The latter authors list 38 Wolf-Rayet stars as probable members of 24 associations. (These are not limited to the 3-kpc distance from the Sun.) The most luminous stars in these associations are blue supergiants of type O, and practically no red supergiants are found. Evolution tracks (Chapter 5) for these clusters imply initial masses for most Wolf-Rayet stars of 50 $M_\odot$, under the conventional assumption that these objects are the first to evolve from the main sequence, and star birth is reasonably coeval. The anticorrelation between the presence of Wolf-Rayet stars and red supergiants, in a well-studied sample, is noteworthy because it implies that most Wolf-Rayet objects never passed through a very cool surface temperature/large radius phase. Humphreys and Davidson (1979) have already called attention to the absence of red supergiants brighter than some bolometric limit. This corresponds to some value between 35 and 50 $M_\odot$, depending on the evolution tracks. The different evolution paths to the helium-burning Wolf-Rayet phase, or the normal red supergiant, are then clear (Chapter 5).

2. z-Distribution of Wolf-Rayet Stars. Figure 2-10 gives the z-distribution of Wolf-Rayet stars within 3 kpc, using the M_v calibration of Table 3-3 (Chapter 3). A comparison with Figure 2-8 indicates common features: a similar scale height, a warped mean z with galactic longitude, and two stars anomalously far from the plane. These stars are HD 50896 and HD 115473. The former is a well-known "WR + collapsed star" *candidate* (Firmani et al., 1980; Moffat, 1982: see Section VI). However, its z-distance in Figure 2-10 was predicted on adopting the mean M_v for a WN5 star of $-4^{\mathrm{m}}.8$. If this object is rather associated with Cr 121 (see Conti et al., 1983a, and Lundström and Stenholm, 1979), its distance is instead 0.7 kpc from us, its z-distance thereby 120 pc, and it is not anomalously placed. HD 115473 appears to be far away from the plane with no other known anomalies. If its distance is correct, it is presumably a descendant from a parent O star with a similar peculiar position.

3. Wolf-Rayet Stars in the Large Magellanic Cloud. Breysacher (1981) lists 101 Wolf-Rayet known stars in the LMC. Four more have been fortuitously discovered since (Conti and Garmany, 1983; Cowley et al., 1984; Westerlund et al., 1985). This listing is believed to be reasonably complete because it is based on recent objective prism surveys specifically for Wolf-Rayet stars which go to sufficiently faint limiting magnitudes (Azzopardi and Breysacher, 1979, 1980). Wolf-Rayet stars in very crowded regions, or with relatively weak emission lines, may have thus far escaped detection. Azzopardi and Breysacher suggest that a complete census might reveal another 40 or so Wolf-Rayet stars.

The Wolf-Rayet star distribution in the LMC, projected on the sky, is given in Figure 2-11. (Note that the linear scale is similar to the O and Wolf-Rayet galactic distributions in

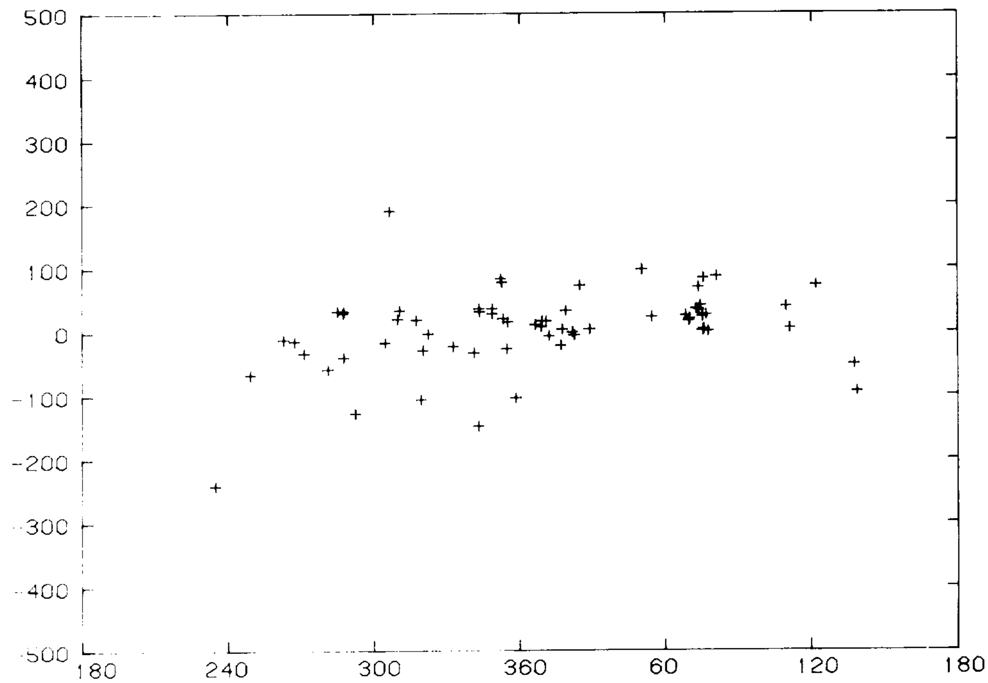


Figure 2-10. Galactic latitude distribution (in terms of z-distance) for Wolf-Rayet stars (plus signs) of Figure 2-9.

Figures 2-7 and 2-9.) If we compare this Wolf-Rayet distribution with the Lucke-Hodge associations, we see a high correlation; many Wolf-Rayet stars are found within the association boundaries, but a large concentration exists near the giant H II region of 30 Dor. There is no obvious concentration to the LMC bar. The Wolf-Rayet stars are also associated with many of the H II regions of the LMC—the Hα emission objects (Davis et al., 1976). Garmany (1984) has thus argued that these correspondences indicate the intimate nature between massive stars and Wolf-Rayet objects.

Most of the Wolf-Rayet stars in the LMC are WN types, and the late WC stars are not found (Smith, 1968a; Vanbeveren and Conti, 1980; Breysacher, 1981). The WN/WC ratio, given the current identification lists, is 4.5, rather different from that of the solar vicinity. According to Melnick (1982), Conti (1982b), and Phillips (1982), there are many more WN stars than WC types in the 30 Dor region; nearly all those within the undifferentiated box in Figure 2-11 are WN types. Aside from the 30 Dor region, the area distribution of Wolf-Rayet stars over the LMC is similar to that of the solar vicinity (compare Figures 2-11 and 2-9). The concentration of Wolf-Rayet stars near 30 Dor is considerably higher than that in the solar vicinity.

The central object of 30 Dor, R136a, was found to have Wolf-Rayet-like features, both in the optical (Melnick, 1982; Ebbets and Conti, 1982), the infrared (Vreux et al., 1982), and the FUV (Savage et al., 1983; Feitzinger et al., 1983). Although this object was initially believed to be a supermassive star ($M \gtrsim 10^3 M_\odot$) according to Cassinelli et al. (1981), it is now believed to be composed of a compact clump of several very luminous stars with masses of a few $10^2 M_\odot$ (Chu et al., 1984; Chu and Daod, 1984; Moffat et al., 1985). At least some of these stars have Wolf-Rayet-like spectra—nitrogen and He II λ4686 emission lines. Such massive stars may evolve homologously, thus showing enhanced nitrogen and helium even while still burning hydrogen in the core (e.g., Maeder, 1980).

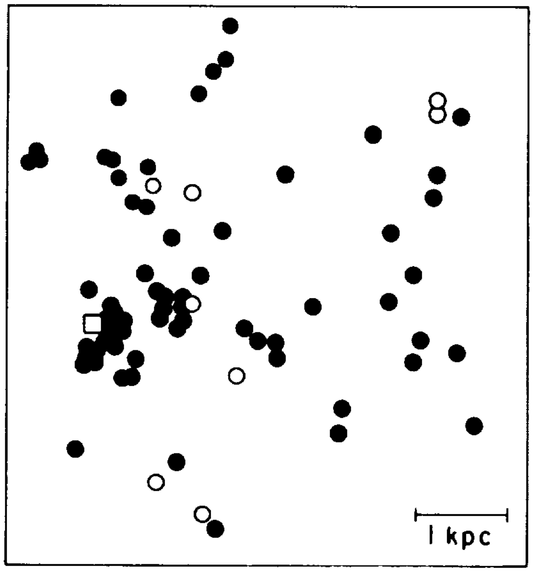


Figure 2-11. Distribution in the sky of known Wolf-Rayet stars in the LMC: open circles denote WC stars; filled circles denote WN stars. The box represents the central region of 30 Dor in which 20 Wolf-Rayet stars are found.

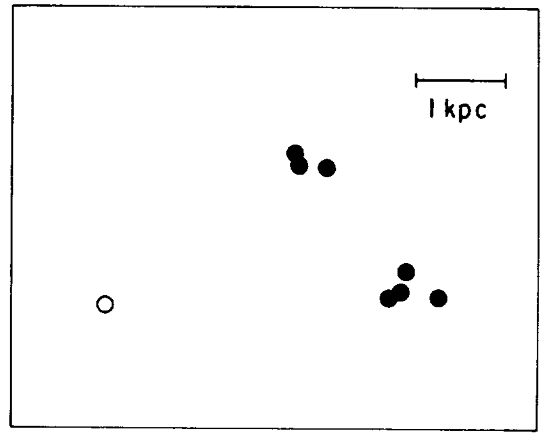


Figure 2-12. Distribution in the sky of known Wolf-Rayet stars in the SMC: open circle denotes WC star; filled circles denote WN stars.

4. Wolf-Rayet Stars in the Small Magellanic Cloud. Azzopardi and Breysacher (1979) list eight Wolf-Rayet stars in the SMC. This is believed to be complete, similar to the LMC because objective prism surveys to sufficiently faint limiting magnitudes have been made. Unlike the LMC, the SMC does not have extensive associations with crowded fields, nor are there many H II regions; therefore, it appears unlikely that any appreciable number of Wolf-Rayet stars have been missed.

One or possibly two of the listed Wolf-Rayet stars are probably Of types (Massey, 1985; Garmany and Massey, 1984). However, all eight are shown in Figure 2-12, which gives the distribution in the SMC plotted on the plane of the sky. (The scale is similar to that of the previous figures.) There is a similarity to the Davis et al. (1976) H II regions (Garmany, 1984).

All except one of the SMC Wolf-Rayet stars are WN types. Thus, the WN/WC ratio is 7 (or 6), following Garmany and Massey (1984). No late WN types are found, and the one WC star is of very early type (Azzopardi and Breysacher, 1979) and more properly called WO (Barlow and Hummer, 1982). On the basis of three "points," it was suggested that the progression in WN/WC ratio from the solar vicinity through the LMC to the SMC might be related to the initial metal abundance (Smith, 1968b; Vanbeveren and Conti, 1980). As we will soon see, this does not hold when other galaxies of the local group are considered (Massey and Conti, 1983b; Armandroff and Massey, 1985).

The area distribution of Wolf-Rayet stars in the SMC is very different from the average LMC area or the solar vicinity (compare to Figures 2-9 and 2-11). Thus, the SMC is deficient per unit area in Wolf-Rayet stars compared to the LMC and our region of the Galaxy.

5. Wolf-Rayet Stars in M33. The first Wolf-Rayet stars to be found beyond the Magellanic Clouds were identified in M33 by Wray and Corso (1972). They studied two fields in this galaxy with direct-image plates, using intermediate-width filters (~100 Å) centered on λ4670 (the λ4650 C III and λ4686 He II features) and λ4450 (a relatively line-free continuum). They found 25 Wolf-Rayet candidates which were brighter "on-line" than "off-line." Two of these candidates were confirmed spectroscopically by Lynds (Wray and Corso), thus suggesting the reality of the Wolf-Rayet stars in the survey. One Wray and Corso object was later found to be a supernova (SN) remnant (D'Odorico et al., 1978), but nearly all of the rest were confirmed to be Wolf-Rayet stars, mostly of WC type, by Wampler (1982), who also found another object from the Wray-Corso finding charts. An "accidental" discovery of an additional Wolf-Rayet star in IC 132 was announced by Boksenberg et al. (1977).

Bohannan et al. (1985) undertook a grism survey of the entire field of M33 in an attempt to discover additional Wolf-Rayet stars. A few Wray and Corso (1972) objects were found, and a handful of other Wolf-Rayet candidates were detected, but the insufficiently deep plate limit and general crowding of images precluded this survey as a useful method. Spectroscopy of the few Wolf-Rayet candidates confirmed their nature (Conti and Massey, 1981b). Serendipitous discovery of the Wolf-Rayet "features" at λ4686 among some very luminous central exciting objects in the giant H II regions was announced by Conti and Massey and, independently, by D'Odorico and Rosa (1981). These luminous exciting objects, with $M_v \sim -8^m$ to -9^m are undoubtedly morphologically similar to the central objects in 30 Dor.

Massey and Conti (1983c) surveyed the field of M33 with a set of deep interference filter plates. The narrowband filters were similar to those used by Wray and Corso (1972). Many new Wolf-Rayet candidates were found, some in the Wray and Corso fields. A number were subsequently confirmed by spectroscopy, leading to a total of 95 known in M33. This list is still incomplete, but probably by less than a factor of 2, as some candidates remain to be confirmed, and "optimized" filters, which were later utilized on prime-focus charge-coupled-device (CCD) frames, led to the discovery of a few more Wolf-Rayet candidates in M33 (Armandroff and Massey, 1985).

From the 95 known Wolf-Rayet stars in M33, Massey and Conti (1983b) found that most were located in the cataloged OB associations of Humphreys and Sandage (1980). A few were found in interarm regions, and not all associations contained Wolf-Rayet stars. Wolf-Rayet and luminous OB stars are again found to be intimately related. In M33, the overall ratio of WN/WC stars is 1, as it is in the solar neighborhood, yet the O/H abundance is overall considerably lower according to Kwitter and Aller (1981). The WN/WC ratio changes with galactocentric distance in M33, ranging from 0.5 near the center to 5 near the outer parts. This change is qualitatively similar to that which I have noted previously in our Galaxy, but does not follow quantitatively. Massey and Conti pointed out that, in the nucleus of M33, the O/H abundance is similar to that of the LMC (according to Kwitter and Aller), yet as we have seen, the WN/WC ratio is 4.5 in the latter and 0.5 in the former. The changing ratio of WN/WC stars in M33 appears to be giving clues to the massive-star progenitors, but whether this is related to the initial masses or the initial composition is not yet clear. Figure 2-13 shows the distribution of Wolf-Rayet stars in M33 at the same linear scale as that of the previous figures.

All of the WC stars in M33 appear to be early types (WCE), as do most of the WN's, according to Massey and Conti (1983b). This

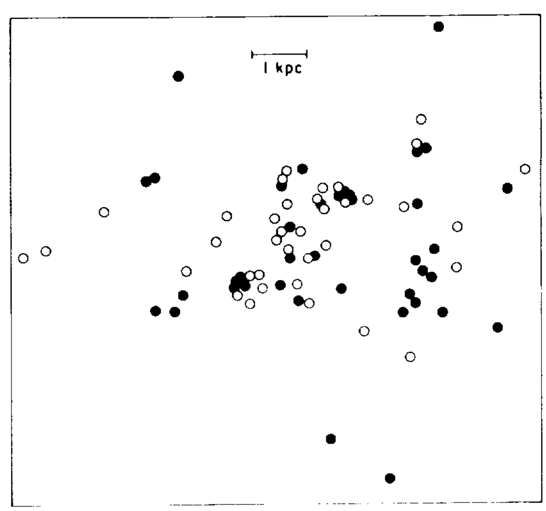


Figure 2-13. Distribution of Wolf-Rayet stars in M33 projected into the plane of that galaxy: open circles denote WC stars; filled circles denote WN stars.

is similar to the LMC for the WCE objects, but not for the WNs, since all subtypes are found. This dichotomy is not yet understood.

D'Odorico and Benvenuti (1983) discuss the properties of the luminous Wolf-Rayet object in the core of the giant H II region, IC 132. Rosa and D'Odorico (1982) investigated the spectra of the knots in NGC 604, the most energetic H II region of M33, and confirm the presence of numerous luminous objects with Wolf-Rayet emission-line features. These spectral appearances are morphologically similar to R136a. According to Massey and Hutchings (1983), this morphological similarity extends to the ultraviolet.

6. Wolf-Rayet Stars in M31. The first comprehensive search for Wolf-Rayet stars in the giant spiral galaxy was carried out by Moffat and Shara (1983). Direct-imaging plates using Wolf-Rayet/continuum filters were obtained, covering two-thirds of the major luminous portion of M31 (but not including Baade's field IV or other regions at this distance). They found 21 Wolf-Rayet candidates, and their later spectroscopy confirmed 17 of them, mostly of WC type. Most of the Wolf-Rayet stars lie in the ring of prominent star formation 7 to 12 kpc from the center of M31.

Moffat and Shara (1983) realized that they may have missed some weak-line WN stars, but nevertheless point out that the total number, even accounting for the area incompleteness, appears to be low compared to our Galaxy, a result they attribute to a low star-formation rate. However, Massey et al. (1986) report the detection of 24 new Wolf-Rayet candidates in four associations of M31. Their photometry was based on CCD imagery and new Wolf-Rayet filters (see the following). The Wolf-Rayet population of M31 has not yet been completely sampled, and all statistical arguments are rather premature.

7. Wolf-Rayet Stars in NGC 6822 and IC 1613. Moffat and Shara (1983), complementary to their M31 work, also obtained plates of NGC 6822 with an on-line Wolf-Rayet filter and a broadband "B." They found no Wolf-Rayet stars to a claimed magnitude limit of $M_B \sim -3^m$, although this "may be slightly too optimistic," and weak-lined WN stars might have been missed. Curiously, almost simultaneously with this paper, Westerlund et al. (1983) announced the discovery of a relatively bright strong-lined WN star in NGC 6822. The serendipitous discovery of a WC star in IC 1613 was announced by D'Odorico and Rosa (1982) and by Davidson and Kinman (1982).

A deep survey of both of these Magellanic-type irregular galaxies for Wolf-Rayet stars has recently been reported by Armandroff and Massey (1985). Their technique involved CCD imaging, enabling them to go fainter in less time, thus increasing the photometric accuracy. More importantly, they used "optimized" Wolf-Rayet filters. These 50 Å wide filters were devised by Massey to isolate separately the λ4650 C III feature from the λ4686 He II one. The former is strong in WC stars; the latter is strong in WN stars. A separate continuum filter of similar width was also used. Figure 2-14, taken from their paper, indicates the relative intensity of the Wolf-Rayet filters. Figure 2-15, also taken from Armandroff and Massey, shows the separation of known galactic Wolf-Rayet stars convolving the filter bandpasses with absolute spectrophotography. There is a nice separation of Wolf-Rayet from normal stars, and WC from WN stars. There is some slight ambiguity of a couple of weak-lined WN stars with normal ones, and between a few WN and WC stars, but at the 90-percent level, these filters will pick out Wolf-Rayet stars and separate them into WN and WC classes.

The Armandroff and Massey (1985) survey identified seven candidate Wolf-Rayet stars in NGC 6822, all of WN type, and five in IC 1613, three WN and two WC. They easily found the previously known Wolf-Rayet star in each galaxy. The numerous CCD frames do not cover each galaxy completely, but well over two-thirds of the projected area in each case was searched. In addition, Armandroff and

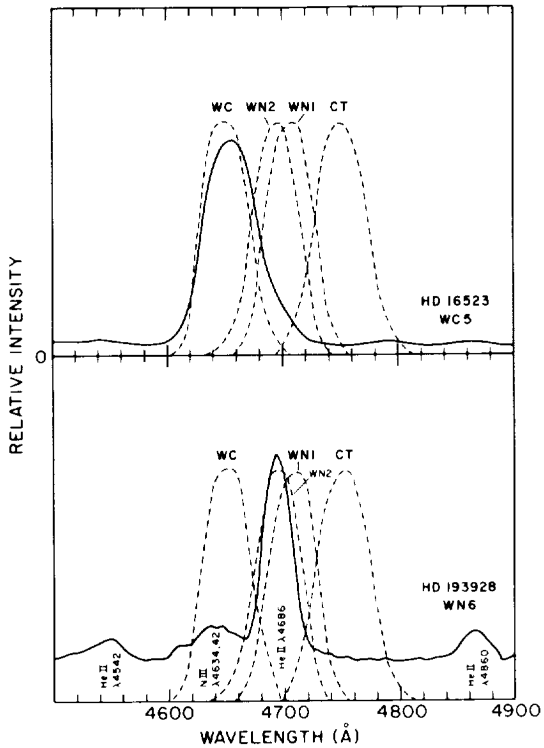


Figure 2-14. "Optimized" Wolf-Rayet detection filters (from Armandroff and Massey, 1985). The WC filter nicely matches the emission-line feature of λ4650 C III in HD 16523; the WN1 filter is a close match to the λ4686 He II emission feature in HD 193928. The combination of these two filters, and a continuum one, isolates Wolf-Rayet objects from normal stars and provides a separation into WN and WC stars.

Massey took two CCD "test" frames in M33, in regions previously surveyed by Massey and Conti (1983b). In addition to the 16 stars already known, they found an additional five Wolf-Rayet candidates, all of WN type. This suggests a possible 30-percent incompleteness in the Massey and Conti survey of Wolf-Rayet stars in M33.

8. Wolf-Rayet Stars Beyond the Local Group. Emission-line features (notably λ4686 He II) have been found in the integrated light of a handful of galaxies beyond the local group. Allen et al. (1976) apparently were the first to call attention to this phenomenon in the dwarf galaxy, He 2–10. Bergeron (1977) noted λ4686 emission in the compact galaxy, IZw18, and suggested a stellar origin. This was followed by observations of Tololo 3 (Kunth and Sargent, 1981) and NGC 5430 (Keel, 1982). Similarly, emission in NGC 6764 and MKN 309 (Osterbrock and Cohen, 1982) and Tololo 89 (Durret et al., 1985) was found. The fact that Wolf-Rayet emission lines are seen at all against the galaxy continuum suggests that an extraordinary number of these objects are implied to be present in the cores of these systems.

There are morphological spectral similarities to the very luminous objects in giant H II regions (D'Odorico and Rosa, 1981; Kunth and Sargent, 1981; Conti and Massey, 1981b; Moffat et al., 1985). The current spatial resolution is insufficient to determine whether normal Wolf-Rayet stars are present or whether many superluminous Wolf-Rayet objects exist in order to explain the spectroscopic observations. With the recently popular concept of "starbursts" in galaxies, it appears likely that Wolf-Rayet stars may be likely consequences of such narrow time events.

Kunth and Joubert (1985) have made a statistical analysis of the Wolf-Rayet emission-line phenomenon in "lazy" galaxies. These are blue systems with emission lines in which it is speculated that star formation goes onward in intermittent short bursts. Spectra of such compact galaxies have been particularly searched for the presence of λ4686. The feature was positively detected in one galaxy (MKN 750) and suspected in 14 more. After allowing for selection effects, Kunth and Joubert conclude that Wolf-Rayet stars are not detected in lazy galaxies with any preferential luminosity or heavy element content. In my view, detection of Wolf-Rayet stars against a galactic background will reveal only the most extreme cases of their presence, and much more work in the area needs to be done and will presumably be in the future. Wolf-Rayet stars may indeed provide key evidence on starburst parameters.

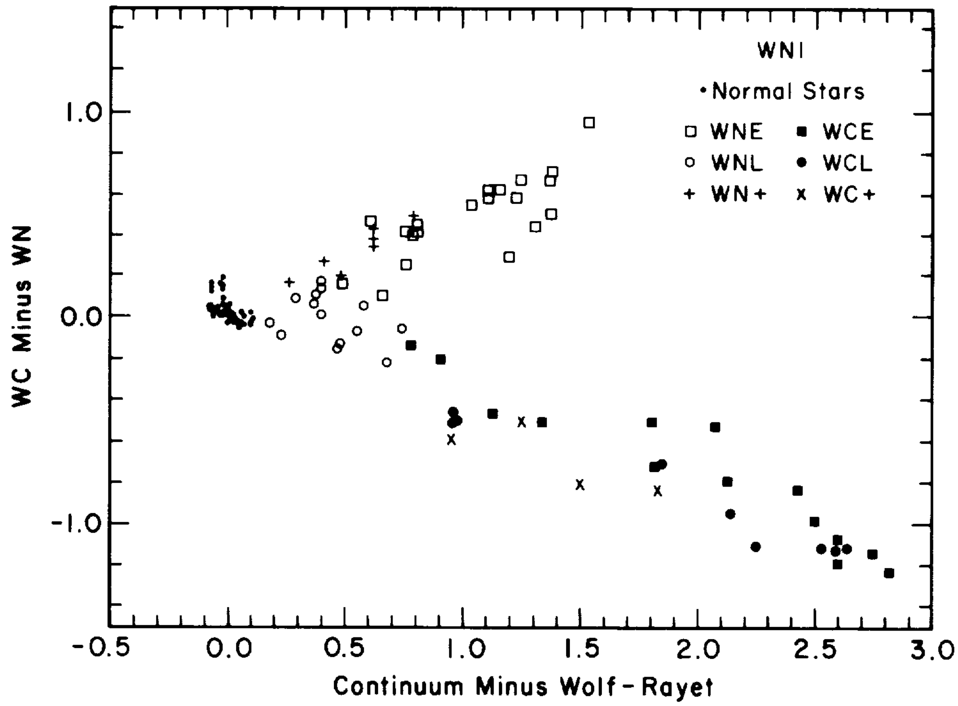


Figure 2-15. "Theoretical" data on galactic Wolf-Rayet stars using the "optimized" filters of Figure 2-14 applied to absolute spectrophotometry (from Armandroff and Massey, 1985). The figure illustrates the separation accomplished between normal and Wolf-Rayet stars and between WN and WC subtypes.

V. KINEMATIC PROPERTIES

A. Rotational Velocities

From coudé spectrograms of 205 O-type stars distributed in both hemispheres, Ebbets and Conti (1977) estimated rotational velocities. Their *V sin i* distribution for main-sequence stars was bimodal with peaks near 100 and 300 km s^{-1}. It was necessary to allow for some macroturbulence because no very sharp-lined supergiants were found. Although their *V sin i* distribution was found to be single-valued, a number of Oe stars were found among the rapidly rotating main-sequence stars (analogous spectra to Be objects), and Oef stars (disk-like emission profiles *and* λ 4686 He II) were found among rapidly rotating luminous stars. The rapidly rotating main-sequence O stars have a velocity distribution similar to that of Be stars (e.g., the high-velocity bimodal peak). Many of them have shown emission sporadically (e.g., ζ Oph). Ebbets and Conti conjecture that members of this group will each show emission at some time. The origin of the bimodal *V sin i* distribution for O stars (and B and Be stars) is not really understood.

Ebbets (1978, 1979) investigated in more detail the question of macroturbulence in O-type stars. He was able to separate the rotational and turbulent contributions by a judicious application of Fourier transform techniques to the line profiles. He found macroturbulent velocities of 30 km s^{-1} in supergiants but little in main-sequence stars. This feature might well be related to the stronger winds in the more luminous stars (Chapter 4)—the broadening of photospheric features a consequence of flow already being important at that level.

Measurements of rotational velocities for Wolf-Rayet stars are essentially nonexistent. Practically no photospheric absorption lines are visible, and the line-broadening in the emission lines is clearly dominated by wind effects. Could Wolf-Rayet stars be rapid rotators? I doubt it for the following reason: HD 50896, one of the broadest lined WN stars, has been found to be a spectrophotometric variable with a period of 3.763 days (Firmani et al., 1980). McLean (1980) has found variable polarization which fits this period; Cherepashchuk (1981) found the same period in his spectrophotometric data. An immediate implication is that this is a rotation period of a nonspherically symmetric object. Such a rotation is *relatively slow*. This star may be asymmetric because of an accompanying collapsed companion (Firmani et al., 1980) although this interpretation is not without its faults (Vreux, 1985b). In any case, our best evidence for rotation in a broad-lined Wolf-Rayet star is that it is relatively slow.

B. Space Velocities

Since O stars are rare in the solar vicinity, they are found, on the average, far from the Sun. Proper motion data are scarce and not always of high accuracy; Stone (1978) has presented some proper motion measurements to which I will return later. Nearly all kinematic data have thus depended on radial velocities. The systematic effects of the Sun's reflex motion and the differential galactic rotation (the Oort constants) must be accounted for in any determination of kinematic velocities. Velocities with systematic effects removed are called "peculiar" velocities.

Most O-type stars are found to have radial velocities consistent with their extreme Population I nature (Conti et al., 1977b); that is, purely circular orbits closely confined to the galactic plane. A net blue-shift can arise if the stellar wind is sufficiently strong that the Balmer lines are formed in the wind (Conti et al., 1977a; Conti et al., 1977b; Bohannan and Garmany, 1978), but this is only an atmospheric effect. Radial velocities from mainly single coudé spectrograms of over 200 O stars were discussed by Conti et al. (1977b). When potential binaries (variable velocities) and outflow phenomena were considered, only a few candidate O stars remained with "high velocity."

As stated previously, some O stars are well above or below the galactic plane. A few of these stars, and some others, also have high velocities or peculiar velocities greater than 30 km s^{-1}. A recent summary of the kinematics of these high-velocity O-type stars is presented by Gies (1985) in his Ph.D. thesis, and I will follow his arguments closely.

The "peculiar" component of the radial velocities displays a gaussian distribution with a standard deviation of 10 to 16 km s^{-1} (Feast and Shuttleworth, 1965; Balona and Feast, 1974; Lequeux, 1979). *Runaway* O stars (a term first coined by Blaauw (1961)) are defined as those having peculiar velocities of more than 3 sigma from the gaussian distribution, or 30 to 35 km s^{-1}. Blaauw had used the term to describe the motions of AE Aur and μ Col, which had determined space motions directed away from the Orion nebulae (Blaauw and Morgan, 1953). These stars were clearly related to this OB association. The term "runaway stars" has come to be applied to all objects with appreciable motion, even if the relation to a parent association is not clear. Blaauw had suggested that such stars were ejected in the breakup of binary systems in supernova explosions by their erstwhile companions. It has been commonly believed that the origin of all runaway stars is the "Blaauw" mechanism worked out by Boersma (1961). (See reviews by van den Heuvel (1978, 1985).) In this generalized scenario, the recoil produced by the SN explosion to one component of a binary pair imparts a substantial velocity to the system, which remains bound. In later evolutionary stages, the initial secondary (now the most massive star) transfers matter to the compact companion (the initial primary), leading to the well-known massive X-ray binaries.

Alternatively, Carrasco and associates (Carrasco and Creze, 1978; Carrasco et al., 1980, 1982) have proposed that the runaway O stars are UV bright stars of low mass which have otherwise been identified in globular clusters.

They imagine that these field stars are horizontal branch Population II objects whose spectra mimic O stars of Population I. That such stars could exist is suggested by Ebbets and Savage's (1982) published IUE and optical work on RWT 152, an O star far from the galactic plane and probably of low mass. Carrasco et al. note that the rotational velocities of O stars appear to be bimodal (Ebbets and Conti, 1977) and suggest that the higher velocity stars are the low-mass analogs. Walborn (1983b) debunks this idea by showing that nearly all the "low-mass" high-velocity O stars in the Carrasco et al. paper are well-known members of clusters and have variable velocity or wind outflow in the absorption lines in all but one case—HD 93521. This star is well out of the plane at $b = +62°$ and can be regarded as a genuine Population II analog (along with RWT 152). The estimated fraction of O-star runaways ranges from 7 percent (Conti et al., 1977b) to 50 percent (Stone, 1979). Carrasco et al. (1980) also believed this fraction to be substantial.

In his thesis, Gies (1985) discusses all the data, studies the binary frequency of a group of runaways, suggests that the two generalized explanations listed above are incorrect, and argues that a third alternative (namely, close three-body encounters) is probably the correct one. Strong gravitational interactions are expected in young clusters of OB stars (Poveda et al., 1967; Allen and Poveda, 1968; Van Albada, 1968). Numerical stimulation studies of close three-body encounters (Hut, 1984a, 1984b) and binary-binary interactions (Mikkola, 1983a, 1983b) indicate that high-velocity objects can readily be produced. Furthermore, the resultant ejection velocities are highly dependent on the stellar masses, the process being more efficient with increasing mass.

Gies (1985) points out that the observed binary frequency of runaway OB stars places important constraints on the models for their origin. In the old disk-population model, the runaway stars are presumed to be descendants of red-giant stars. Both long- and short-period binaries are detected as nuclei of planetary nebulae (NPN), which are almost certainly red-giant descendants. According to Acker (1984), the binary fraction of NPN is near 40 percent, similar to the cluster and field Population I O-type star fraction of 31 percent (Garmany et al., 1980). The supernova ejection model suggests that runaway stars will be either single or single-lined spectroscopic bodies with low-mass compact companions. No visual companions would be expected. The ejection hypothesis predicts that most runaways will be single stars.

Gies (1985) took a set of 36 OB program stars which have been listed as potential runaway objects. He then extensively observed these stars for a 2-year period to establish their velocities and binary nature. His data are based on 12 to 15 Å/mm spectrograms obtained at the David Dunlap Observatory, which enabled him to determine velocities to a few km s^{-1}. This is by far the most accurate systematic study of O-star velocities yet attempted. Because Gies is limited to northern hemisphere stars, the survey is not complete over the sky, but the results are significant. He finds that, of the 36 runaway candidates, 15 are true runaways and 5 more are probable members of this set, based on their distance from the galactic plane and large proper motion. The other 16 stars are not kinematically peculiar, but have anomalous velocities because they are previously unrecognized binaries (five stars) or have pulsational instabilities (possibly nonradial—seven stars) or insufficient data (four stars). Of the 20 runaway stars, only two are binaries, and both of these are double-lined systems. Thus, Gies points out that the binary fraction of the well-established runaway stars is very low compared to the normal Population I OB stars (31 percent according to Garmany et al. (1980)) and the NPN (40 percent according to Acker (1984)). Gies furthermore points out that none of the confirmed runaways are visual binaries, although seven of the other stars which were initial runaway candidates have these companions. Also, only one of the known X-ray binaries has potential runaway characteristics.

Thus, Gies (1985) is led to the important conclusion that the runaway OB stars are probably ejected from clusters and associations in which close interactions, or binary-binary interactions, have occurred. This is because the observed low binary frequency of runaways is predicted by the ejection model and is inconsistent with the UV bright Population II and the supernova ejecta model predictions.

O stars have lifetimes of a few million years, and the typical velocities of 10 km s^{-1} will carry them only a few tens of parsecs from their birthplaces. However, stars with appreciable space velocities, such as these runaway objects, could be found much farther away. For peculiar velocities of 50 km s^{-1}, distances of 200 pc could be reached if the close interaction ejection occurs *early* in the lifetime of a massive star, as seems likely. The supernova ejection scenario always suffered from this defect; the event occurred late in the evolution of the system, thus giving it little chance to move far from its starting point. It appears likely that not only runaway stars but also OB stars at appreciable distances above or below the galactic plane can be found there by close interaction events near the time the stars formed. Although not all of the anomalously placed objects currently have high velocities, it should be remembered that their present-day motions may have been influenced by the gravitational forces perpendicular to the galactic plane.

What of the proper motions? In a series of papers, Stone (1978, 1979, 1981, 1982) presents measurements for a set of O stars and, similarly to Carrasco and associates, purports to demonstrate that a substantial subset of "well-known" O stars have peculiar space velocities. He argues that his high-velocity objects are ejecta from SN explosions and should therefore be identified with that scenario. My impression on going over the papers is that Stone is greatly overinterpreting his data. There is little correspondence between the high-velocity stars from the proper motion measurements and those of Carrasco and associates. Similarly, several are well-known cluster members and/or have known variable velocities. In particular, in Stone (1979), a careful inspection of the 62 luminous stars with space velocities reveals significant (6 σ) results only for AE Aur, ξ Per, HD 157857, and 9 Sge. The first two are well-known anomalous-velocity stars. The other high-velocity stars in Stone's tables have smaller significance if the proper motion uncertainties are larger than his formal estimates. My opinion is that systematic effects may be important in the proper motion measurements, an issue not adequately addressed by Stone (cf. controversy of the Hyades distance from the moving cluster method—and these O-type stars are much farther away).

Tobin and Kilkenny (1981) have studied the properties of OB stars (mostly type B) far from the galactic plane. They compare 16 stars with $z > 1.5$ kpc, with 14 in or near the plane. There were no $V \sin i$ differences, no gravity differences (from Hγ profiles), and no other anomalies. These high-latitude OB stars are apparently normal Population I stars; HD 93521 remains a notable exception and is probably Population II. Keenan et al. (1982) analyze some other "halo" OB stars and find normal He, C, N, O, and calcium.

As far as Wolf-Rayet stars are concerned, their kinematic velocities are quite uncertain. This is partly because the emission-line widths make measurement difficult, but also because there are velocity gradients among these lines which are formed in the outflowing wind. Furthermore, variability of line profiles may be a common characteristic of Wolf-Rayet stars (Vreux, 1985b, and this complicates interpretation. Torres and Conti (1984) have discussed some of the problems inherent in determining the systematic velocities of Wolf-Rayet stars. Conversion to peculiar motions is further complicated by less certain distances, thus compounding the correction for galactic rotation.

VI. BINARY FREQUENCY

A. O-Type Stars

Garmany et al. (1980) have studied the binary frequency of a sample of 67 O-type stars which are fairly bright and thus amenable to high-dispersion spectroscopic analysis. They find that 36 ± 7 percent of the stars are binaries, mostly with mass ratios less than 2.5. The sample size and velocity detection procedure suggests that very few undetected binaries exist in their sample. Although O stars typically have relatively few lines, which are often quite broad compared to, say, F-type stars, their large masses make it relatively easy to detect all binary systems of short period since the velocity ranges will be necessarily substantial. Companions with orbital periods of years will, of course, not be easily found.

The actual tabular listing of the known O-type binaries is contained in Chapter 1 of this volume (Table 1-25). A few additional systems have been investigated since this compilation. In particular, Niemela (1986) has found three new O-type binaries in the Large Magellanic Cloud, all of them double-lined systems.

Do any O-type systems contain collapsed (neutron star) companions? Certainly the known X-ray binary systems, such as HD 153919 = 3U1700-37, are examples of this. But are there X-ray "quiet" systems with such companions? Such theoretical expectations have been proposed. (See for example, van den Heuvel, 1978.) The proposed generalized scenario is as follows: the known X-ray binary systems must have begun as O-type binary systems. The initially more massive component evolves first and becomes a Wolf-Rayet star with an O-type secondary companion. Such systems have been observed. The Wolf-Rayet star then completes its evolution first, presumably by a supernova explosion to a neutron star or black hole. The system remains bound and results in an O star (the initial secondary) with a collapsed companion. Van den Heuvel suggested that many such "quiet" systems exist at an evolutionary stage prior to the advent of substantial wind activity and/or expansion to the Roche lobe. At that point, substantial matter would flow to the collapsed companion, and the system would be subsequently identified as a known strong X-ray binary system. Are such hidden and quiet systems found? Garmany et al. (1980) examined this question and suggested several candidate stars with possibly low-velocity amplitude and thus low-mass close companions. No subsequent radial-velocity studies have yet been made to confirm or deny these candidates. For now, the question of "X-ray quiet" collapsed companions for O stars remains open.

B. Wolf-Rayet Stars

Vanbeveren and Conti (1980) have considered the statistics of Wolf-Rayet binaries among a large sample of galactic and Magellanic Cloud Wolf-Rayet stars. Their study concerned the appearance of absorption lines among a number of these objects. In most cases, absorption lines are indicative of an O-type companion, although a few demonstrable exceptions to this assertion can be found (Conti et al., 1979). Vanbeveren and Conti noted that less than 40 percent of the Wolf-Rayet stars in their extensive sample showed evidence of absorption lines, and certainly a few of these are intrinsic to the star itself or are due to a nearby, but not close, companion O-type star. They also argued that most systems with O-type companions to Wolf-Rayet stars would be sufficiently bright that the absorption lines would be visible. Thus, their result was that the binary frequency of Wolf-Rayet stars with O-type companions was near the observed 40-percent limit. Massey et al. (1981) also address the question of absorption-line Wolf-Rayet systems and, in a companion study, find a binary frequency of this same order.

The listing of the known Wolf-Rayet binary systems is contained in Table 1-28 of Chapter 1 of this volume (taken mostly from Massey, 1981). Since then, information concerning a few additional Wolf-Rayet binary systems containing O-type companions has been discussed, mainly by Niemela and associates (e.g., Niemela and Moffat, 1982; Niemela et al., 1986). The binary frequency of O and Wolf-Rayet systems appears to be similar and near 40 percent for each. It is hard to imagine that the binary nature of some of the Wolf-Rayet systems is at all related to their anomalous spectra because the known single stars have similar properties.

Do any Wolf-Rayet stars exist in systems with collapsed close companions? Again, these are expected theoretically. Following the van den Heuvel binary scenario outlined previously, we would expect that the known X-ray binary systems, containing an O star and a collapsed companion, should evolve further into a Wolf-Rayet object and the same collapsed companion as evolution proceeds. Simple arguments (e.g., as outlined by Vanbeveren and Conti, 1980) suggest that about as many Wolf-Rayet stars with collapsed companions as with O-type stars should exist. However, no known strong X-ray binary systems contain Wolf-Rayet stars as companions. Could there be X-ray "quiet" Wolf-Rayet stars, plus collapsed companions? The answer to this question is currently somewhat controversial.

On the one hand, Moffat and associates have photometrically and spectroscopically studied a sample of Wolf-Rayet stars (generally without absorption lines) and found apparent periodic variability in a number of them. (Lists and extensive references are contained in Moffat (1982).) The periods seem to be a few days in many cases; the same period is often found in the photometric and spectroscopic data, and the mass functions are small, thus suggesting that the masses of the companion objects are also low. Some systems have more compelling data than others. For example, HD 197406 (Drissen et al., 1986) has a well-determined radial-velocity curve which is unquestionably due to a companion object. Polarimetric observations suggest a nonspherically symmetric shaped Wolf-Rayet star because periodic variability with the same period as the orbit is found. These authors contend that a black-hole companion to HD 197406 is the most likely explanation for the data.

On the other hand, Vreux (1985b) has questioned the uniqueness of the periodicities found by Moffat and associates in some other stars. In one case, that of the system HD 192163, Vreux et al. (1985b) found that a period of a few hours fit the radial-velocity data they obtained. The longer period of a few days previously published could be interpreted as an alias, or a rotation period. These authors suggested that, in some cases, the true periodicities in the emission-line shapes were shorter than a day (often the data published could not unambiguously differentiate between such choices) and were caused by nonradial pulsations in the Wolf-Rayet stars themselves and not due at all to a companion.

That subtle variations are seen in Wolf-Rayet-star emission lines in the optical seems to be established beyond doubt. Aside from the case of HD 197406, these are difficult to measure, typically being of the order of 1 percent of the linewidths and requiring the highest signal-to-noise data. Variable absorption features are sometimes found superimposed on emission lines (e.g., Firmani et al., 1980). Although there is some information on day-to-day variability in these objects, as yet not much effort has been paid to much shorter period variability. I understand that considerable work is currently in progress, both with visual data and with the IUE satellite. My hunch is that the variability we observe will, in many cases, turn out to be related more to nonradial pulsations or some other instabilities in the atmospheres and winds of Wolf-Rayet stars rather than in the existence of collapsed companions in orbit around them. I believe that the periods of a few days which seem to be well established in some cases (e.g., HD 50896) could well be the fundamental stellar rotational period.

Let me end this section by saying that these issues are far from settled and will require substantially more data than we have available at present. The binary scenario would suggest that Wolf-Rayet stars with collapsed companions should exist; if they are in fact not found, reasons will have to be found to explain this. If those Wolf-Rayet objects identified by Moffat and associates are indeed such systems, one will alternatively have to explain why they are not X-ray sources (e.g., Vanbeveren et al., 1982).

3

INTRINSIC STELLAR PARAMETERS

I. ABSOLUTE VISUAL MAGNITUDES

A. O-Type Stars

The relationship between spectral type and absolute visual magnitude, M_v, for O stars is based on cluster membership, in which the cluster distance has been determined through the B-star absolute magnitude calibration. This ultimately depends on cluster main-sequence fittings, back to the Scorpio-Centauris association distance (Blaauw, 1961). "Modern" estimates of these M_v magnitudes were made by Conti and Alschuler (1971) and Walborn (1972). Although the spectroscopic criteria were not identical, they were similar, and the calibrations were not very different.

Conti et al. (1983a) reexamined the O-star calibration, based on 248 O stars in 63 clusters; this was a significant increase in the number available 10 years earlier. The bulk of the data were obtained from Humphreys' (1978) studies of the brightest stars in our Galaxy. This M_v-spectral-type relation is given in Table 3-1. This compilation contains the most recent rediscussion of the earliest O types by Walborn (1982b). The standard error, σ, is the statistical uncertainty of a single observation, not of the mean; n is the number of stars at that spectral type. For most types, σ is of the order of $0^m.5$ and is therefore a basic uncertainty in determining an M_v from a single spectral type alone. Note that the listed mean values of M_v do not change smoothly with spectral type, although the dispersion here ($\approx 0^m.3$) is consistent with the statistical uncertainties of the M_v *means* ($\sigma \div n^{1/2}$).

According to Conti et al. (1983a), the scatter in the M_v-spectral-type relation is intrinsic to the stars themselves. Table 3-1 lists the best available M_v calibration for O spectral types. Although there is a $\sim 1^m$ distinction between the adjacent luminosity classes in the late-type O stars, this is halved in the earliest types, and type III is ill defined. Among the earliest types, the luminosity class depends primarily on the appearance of λ4686 He II, which is a wind phenomenon; in the late-type O stars, purely absorption-line information such as the Si IV line strengths are used. These lines depend directly on atmospheric phenomena such as temperature and effective gravity.

B. Wolf-Rayet Stars

For Wolf-Rayet stars, the magnitude calibration for galactic stars can be estimated from cluster membership, as has recently been thoroughly discussed by Lundström and Stenholm (1984b); it has historically also been determined from Wolf-Rayet stars in the Large Magellanic Cloud (LMC) (Smith, 1968b). Conti et al.

PRECEDING PAGE BLANK NOT FILMED

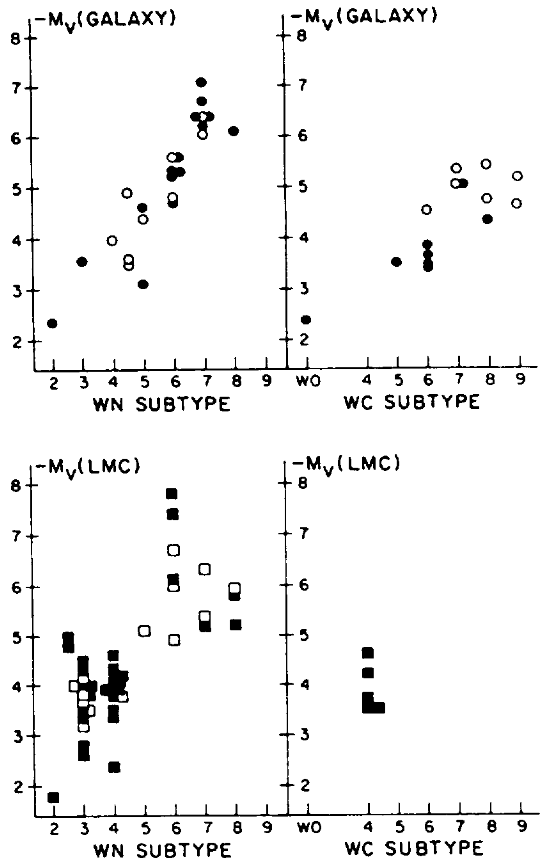


Figure 3-1. Absolute visual magnitudes for WN and WC stars in the Galaxy (upper) and LMC (lower). In the Galaxy, only those stars which have well-established cluster/association membership are indicated (see text). In the LMC, only those stars in which the effects of visual companions have been addressed are shown (see text). The filled symbols have better determined photometry.

stars, but the value of $1^{\mathrm{m}}.0$ for the WN6 σ is disturbing. The WNE stars are fainter than WNL stars (and similarly for WCE and WCL), but there is not a smooth progression. These M_v values for WR stars are the best available data but should not be overinterpreted. The range of M_v from the brightest WN6 star in the LMC (~ -7.8) to the faintest WN2 (~ -1.8) indicates a huge dispersion in this parameter.

Conti et al. (1983a) investigated whether or not the M_v at a given subtype could be related to other spectroscopic properties (for example, the emission line equivalent width, line width, or line intensity). There were not any convincing relations found among the early-type WN stars in which there was a sufficient sample. We can possibly understand this absence of an M_v correlation with emission-line strength to be because the spectra of Wolf-Rayet stars are primarily of their winds, the properties of which are only loosely coupled with their luminosities. However, we are discussing only the M_v here; the M_{bol} may be significantly different, as we will see shortly.

II. ATMOSPHERIC MODELS

A. O-Type Stars

Models are used to numerically describe the run of physical parameters such as the temperature, pressure, and density of electrons and ions, the radiation field, the composition, and, in some cases, the velocity with radius or optical depth. These numerical models are then used to predict observables such as the emergent radiation in both the lines and the continuum. Good models are based on appropriate physics and give quantitative predictions which agree with all of the observable data. Poor models either ignore inconvenient physical principles or substantial parts of the data. Let it be said at the outset that the current "best" models are somewhere between these two extremes. Many observables are reasonably well explained (e.g., the absorption-line spectra of O stars); some quantities (e.g., turbulence) are not well described. Considerable physics are included (e.g., non-LTE in the hydrogen and helium transitions); much is not yet taken into account (e.g., a complete treatment of radiative transfer for all lines). All models are thus incomplete but are useful constructions for gradually advancing our understanding of complicated physical phenomena.

I will limit my discussion here to those modern models used to describe the (assumed to be static) atmospheres of O-type stars. (Models with mass flow are considered in Chapter 4.) In order of increasing complexity

Table 3-1
M_v-Spectral-Type Relation for O Stars*

Type	M_v	σ	n	Type	M_v	σ	n	Type	M_v	σ	n
O3 V	−5.4	0.5	5					O3 I	−6.2	—	2
O4 V	−5.8	0.6	5					O4 I	−6.4	—	3
O5 V	−5.2	0.8	7	O5 III	−6.1	—	8	O5 I	−6.9	—	2
O6 V	−5.1	0.6	11	O6 III	−5.7	0.3	6	O6 I	−6.5	0.6	3
O6.5 V	−5.1	0.8	13					O6.5 I	−6.7	—	1
O7 V	−4.9	0.5	18	O7 III	−5.6	0.6	6	O7 I	−6.5	1.1	7
O7.5 V	−5.0	—	2	O7.5 III	−5.6	0.4	4	O7.5 I	−6.7	—	1
O8 V	−4.6	0.6	19	O8 III	−5.1	0.5	3	O8 I	−6.5	0.9	6
O8.5 V	−4.9	0.9	4	O8.5 III	−5.0	0.5	3	O8.5 I	−5.9	—	2
O9 V	−4.4	0.6	35	O9 III	−5.3	0.7	16	O9 I	−6.1	0.8	8
O9.5 V	−4.0	0.6	17	O9.5 III	−5.1	0.6	9	O9.5 I	−6.0	0.7	22

*Adapted from Conti et al. (1983a).

(1983a) rediscussed the LMC calibration, based on the M_v (this subscript is in the narrowband Smith system) data of Prévot-Burnichon et al. (1981), supplemented by the improved photometry of Massey and Conti (1983a) or spectroscopy (Conti et al., 1983b). I will consider together the recent cluster work of Lundström and Stenholm with the LMC calibration.

The narrowband (*ubvy*) Smith (1968a) system was designed to be relatively line-free, but in strong-lined WC-type stars, this idealized situation is not realized (Massey, 1984). Nevertheless, there are a great number of narrowband measurements of galactic and LMC stars (e.g., Lundström and Stenholm, 1979, 1980; Turner, 1982). Absolute spectrophotometry for all known Wolf-Rayet stars would be useful, and the first data have been published (Massey, 1984; Massey and Conti, 1983a; Lundström and Stenholm, 1984a). The absolute visual magnitudes discussed here, which are on the narrowband *v* system, are thus still uncertain but are the best currently available. In Figure 3-1, I show the individual M_v values for Wolf-Rayet stars in galactic clusters and associations (adapted from Lundström and Stenholm, 1984b) and the LMC (from Conti et al., 1983a). A comparison between the Galaxy and LMC indicates good agreement between the M_v for the early WN subtypes and WC5 stars (late WC types are not known in the LMC). There is surprisingly poor agreement between the M_v of the WNL stars in the Galaxy and LMC; the WN6 stars in the latter are appreciably brighter, and the WN7 and WN8 are appreciably fainter. Frankly, I don't know whether this is "small number statistics" or indicative of some real effect.

Table 3-2 shows the same M_v for WN and WC stars in the Galaxy clusters and in the LMC. The σ is the uncertainty of a single observation, n is the number of stars in a single spectral subtype. In the "all" column, I have simply averaged the Galaxy and LMC M_v values weighted to the number of stars. The σ here are, on the average, $0^m.5$, similar to that for O

Table 3-2
M_v-Spectral-Type Relation for Wolf-Rayet Stars

	Galaxy			LMC			All		
	$-M_v$	σ	n	$-M_v$	σ	n	$-M_v$	σ	n
WN2	2.4	—	1	1.8	—	1	2.1	0.4	2
WN2.5		—		4.9	0.1	2	4.9	0.1	2
WN3	3.6	—	1	3.7	0.5	15	3.7	0.5	16
WN4	4.0	—	1	3.8	0.5	14	3.9	0.5	15
WN4.5	4.0	0.8	3		—		4.0	0.8	3
WN5	4.0	0.8	3	5.1	—	1	4.3	0.8	4
WN6	5.2	0.3	7	6.5	1.0	6	5.8	1.0	13
WN7	6.5	0.3	7	5.6	0.6	3	6.2	0.6	10
WN8	6.1	—	1	5.6	0.4	3	5.8	0.4	4
WC4	—	—	—	3.9	0.5	5	3.9	0.5	5
WC5	3.5	—	1		—		3.5	—	1
WC6	3.8	0.4	5		—		3.8	0.4	5
WC7	5.1	0.2	3		—		5.1	0.2	3
WC8	4.8	0.6	3		—		4.8	0.6	3
WC9	4.8	0.4	2		—		4.8	0.4	2
WO	2.4	—	1		—		2.4	—	1

and sophistication, these are models based on the first modern non-LTE work of Auer and Mihalas (1972), Mihalas (1972), Kudritzki (1976), Kurucz (1979), and Abbott and Hummer (1985). The first three references used codes that included non-LTE calculations (in hydrogen and helium) and plane parallel, radiative equilibrium (RE), and hydrostatic equilibrium (HE) assumptions. These predicted the emergent line and continuum radiation from the stellar atmosphere for a set of models of various effective temperatures and gravities. Kurucz's models included an estimate of line-blanketing, but used LTE. Abbott and Hummer adopted the Auer-Mihalas code and added a crucial feature: incident radiation on the top of the atmosphere backscattered from a stellar wind. As we shall see, this fundamentally alters the emergent flux of hot luminous stars with strong winds. Only one set of models has been attempted thus far by Abbott and Hummer. These results will be considered in due course.

Non-LTE is crucial for understanding the absorption-line spectra of O-type stars and is, in fact, reasonably successful in predicting the line strengths of the hydrogen, He I, and He II lines in most O-type stars (e.g., Auer and Mihalas, 1972; Conti, 1973a). The assumptions of RE and HE appear to be reasonable for O stars insofar as the overall absorption-line and continuum properties in the visual are concerned.

Nearly all non-LTE models have used only hydrogen and helium ions to determine the structure of the atmosphere. The importance of the CNO elements on the continuum opacity has been considered by Husfeld et al. (1984) and found to be minor. The blocking of the outflowing radiation by lines, particularly in the FUV regions, was considered only in an LTE approximation by Kurucz (1979). Anderson (1985) has developed an improved technique to calculate the non-LTE line-transfer problem for other ions. This might make an appreciable difference in the blocking problem when the full process of calculating hundreds or even thousands of lines is completed. This is an important next step in the modeling of O-type atmospheres.

Line profiles offer an additional constraint on the O-star photospheric models. Kudritzki and associates (Kudritzki, 1980; Kudritzki et al., 1983; Simon et al., 1983) have studied high-dispersion photographic spectra to derive line profiles for a number of early O-type stars. Here the data are better than those available previously. The non-LTE models were also extended to lower gravity cases. These authors found good agreement in the predicted and observed line-profile spectrum of Balmer, He I, and He II line cores and wings, particularly in going to a lower gravity, lower temperature model for the O4f star, ζ Pup. I shall return to the actual numbers in the following sections.

Line profiles accurate to 1 percent, substantially better than can be obtained with photographic plates, can be obtained with modern charge-coupled-device (CCD) detectors on high-dispersion spectrographs. Abbott and Hummer (1985) speak of the necessity of constructing model atmosphere line predictions which will meet this accuracy.

Hummer (1982) has demonstrated that the backscattering of photons from a strong stellar wind might be an important boundary condition for hot-star model atmospheres. For example, he called $\bar{\rho}$ the mean fraction of ultraviolet (UV) photons backscattered. Then $(\bar{\rho}/1-\bar{\rho}) = 0.3\ \dot{M}v_\infty\ c/L$ can be used to estimate $\bar{\rho}$. The $\dot{M}v_\infty\ c/L$ term is ~ 1 for typical O stars and ~ 10 for Wolf-Rayet stars with generally accepted values of $\dot{M}$, v_∞, and L. Thus, $\bar{\rho} \sim 0.23$ for O stars and ~ 0.75 for Wolf-Rayet objects, an important effect for the former and a dominant one for the latter.

Abbott and Hummer (1985) have constructed models for ζ Pup which incorporate the effect of wind-blanketing. Basically, the upper boundary condition which specifies no radiation incident down upon the photosphere is replaced by a nonzero incident field. This applied condition is the emergent continuum radiation intensity, multiplied by a wavelength-dependent albedo. This albedo is taken from a Monte Carlo solution of the transfer equation in the stellar wind, incorporating 10^4 UV and visible lines (Abbott and Lucy, 1985). This modeling is only a first step, but the inclusion of the backscattering makes a substantial difference in the line and continuum radiation. The most important result is that the difference between the boundary and effective temperatures is *lessened*; this particularly affects the He I line strengths and thus the relationship between the spectral type and inferred effective temperature *for stars with strong winds*. Excellent agreement between the line-profile predictions and new high-accuracy measurements in ζ Pup is reported by Bohannan et al. (1986). An extension of the technique to a larger range of models and detailed comparisons to more stars is planned.

B. Wolf-Rayet Stars

Because realistic model atmospheres for Wolf-Rayet stars must incorporate the physics implied by the existence of a wind from the beginning, the atmosphere and wind cannot be treated separately and independently. The wind opacity is so great (Chapter 4) that no radiation from the hydrostatic core is seen. All the observed atmosphere is in appreciable motion. The assumptions of plane parallel geometry, hydrostatic equilibrium, and even radiative equilibrium may not be even approximately applicable in real Wolf-Rayet stars. Extended spherical geometry must certainly be used.

Mass-loss rates of $\sim 3 \times 10^{-5}\ M_\odot\ \mathrm{yr}^{-1}$ are found for many Wolf-Rayet stars by a relatively model-independent method which is based on the detection of free-free emission in radio frequencies (Barlow, 1982; Chapter 4).

Radiative-transfer and statistical equilibrium equations must be solved in an expanding atmosphere for an initial estimate of predicted emission-line strengths in Wolf-Rayet stars. A major advance in this type of approach, even though considerable physics are still not included, has recently been provided by Hillier (1983). He obtained new data on infrared (IR) lines of the helium ions in HD 50896, a relatively nearby WN5 star with very broad emission lines, and combined this with existing visible and ultraviolet (UV) measurements. He then attempted to model the observed line profile under *assumptions* of spherically symmetric extended geometry, a monotonic velocity law, HE, RE, and pure helium composition. Hillier did not use the Sobolev (1960) approximation, but attempted exact solutions of radiative and statistical equilibrium equations for the helium ions, following the techniques of Mihalas et al. (1975). He treated the continuum and line-formation problem simultaneously. From a comparison of the models with the line and continuum data, he found that the continuum energy distribution is primarily determined by the wind density, not by the effective temperature. (See Section III.B.) Electron scattering plays a role in the formation of the He II lines. The velocity law is relatively "slow," reaching 60 to 70 percent of its terminal value at 50 stellar radii according to the models. Because the helium ionization balance is not "frozen-in," but shifts from predominantly He^{++} to He^{+} within the wind, the He I and He II lines originate in somewhat different atmospheric regions in this star. A model with $\dot{M} \sim 5 \times 10^{-5}\ M_\odot\ \mathrm{yr}^{-1}$, ("core" radius) $R_c \sim 2.5\ R_\odot$, $L \sim 5 \times 10^4$, and T_e (electron) in the wind 30000 K best fits the observations. Hillier also finds that HD 50896 must be very hydrogen-deficient because no excess contribution to the Balmer lines, *vis a vis* the Pickering He II lines, is seen.

Hillier's (1983) sets of models are a starting approximation to real WN stars. Previous attempts have used only a few visible or UV lines, or the continuum, but none of these together. The next step will be to incorporate the back-scattering of the wind photons on the underlying radiation field, following Abbott and Hummer (1985). Such an undertaking might then be considered a first approximation to a real Wolf-Rayet-star atmosphere and wind.

III. EFFECTIVE TEMPERATURES

Although the concept of an effective temperature is, in principle, simple, its application in the literature is sometimes filled with pitfalls and misapplications. By definition,

$$L \equiv 4\pi R^2 \sigma T_{\mathrm{eff}}^4 \quad , \qquad (3\text{-}1)$$

where L is the bolometric luminosity of the star, R is the radius of the "surface" (defined in Chapter 2), and T_{eff} is the effective temperature. No problems are encountered if the radius from which the continuum radiation emerges is independent of wavelength and is physically at the same "depth" in the star. However, if this radius varies with wavelength over the observed regions, serious misinterpretations may result. In Wolf-Rayet stars, the opacity of the wind is sufficiently high that the radius is observed to be wavelength-dependent (e.g., color-dependent eclipse durations for the WN system, V444 Cyg (Cherepashchuk et al., 1984)). To use Equation (3-1) to derive T_{eff} directly, one must have estimates (measurements) of L and R. The former will depend on model-dependent bolometric corrections since a substantial fraction of the emitted radiation from O and Wolf-Rayet stars lies in the unobservable FUV regions (below 912 Å). Not many data are as yet available below ~1200Å, the short-wavelength limit of the International Ultraviolet Explorer (IUE). The quantity R can be found in principle from direct measurements, using interferometry, or from eclipses (which is discussed in Section V).

A. O-Type Stars

Direct angular diameter measurements of a few O-type stars have been made (e.g., Hanbury-Brown et al. (1974)) for ζ Pup. Incidentally, Kudritzki et al. (1983) estimate that the electron scattering in the wind of ζ Pup leads to a ~2-percent overestimate of the radius, or a ~4-percent uncertainty in T_{eff} by the direct method. However, for most stars, R cannot be measured directly but must be found indirectly.

A more common method for deriving effective temperatures is to construct a series of models that most readily represent observations such as continuum radiation, line spectrum, etc. and then adopt the appropriate interpolated model T_{eff}. Kudritzki and associates have stressed that one must simultaneously solve for the effective temperature and effective surface gravity, g_{eff}, by appropriate models. Abbott and Hummer (1985) point out that the wind strength (via the backscattering) is also critical. Hence, model atmosphere determinations of hot-star continua and spectra are at least a *three-parameter family*. Aside from ζ Pup, only T_{eff} and gravity models have been used thus far in the literature, and I will be constrained by that in what follows.

Note also that astronomers dealing with stellar structure models and evolution like to place their stars in log L, log T_{eff}, (or log g_{eff}, log T_{eff}) diagrams. In doing this, they must adopt a surface boundary condition often simply $P = 0$ (or $\rho = 0$); this sets the radius and therefore determines a T_{eff} which is model-dependent. If a photosphere is thin in radius, this simplification produces no important error, but since some real stars, such as Wolf-Rayet objects, have extended geometry, serious errors of interpretation of stellar effective temperatures can result (de Loore et al., 1982).

Likewise, for observers, the easy data include magnitudes, colors, and spectral types. These must be calibrated, by using atmospheric models, into the stellar structure predictions of L and T_{eff}. This calibration step is only as good as the models. Recognizing that readers will want some discussion and tabulation of T_{eff} for hot luminous stars, I will give here a critical current compilation for the O-type stars. These are based mainly on the theoretical prescriptions listed previously, except that backscattering models have thus far been applied only to ζ Pup. My best current guess as to the effects on other O-type stars is that only in strong wind stars is this phenomenon important, at least as we now understand the situation.

I will not list T_{eff} for Wolf-Rayet stars because I consider them to be uncertain by at least 50 percent. I believe that they range from 30000 K to two to three times that value or more, but this statement is not much of a change in some years (e.g., Conti, 1976). Backwarming from the wind undoubtedly influences the physical situation and has not yet been modeled. A table of T_{eff} for Wolf-Rayet stars is provided in Part I of this volume. Part III also offers alternative interpretations of the state of knowledge of T_{eff} for O and Wolf-Rayet stars.

First, let me list the applicable references from the literature. I will then combine the numbers and estimate a T_{eff}-spectral-type diagram. Conti (1973b, 1975) provided a list of T_{eff} for many measured O stars, based on calibrated spectral types (measured He I/He II line ratios) and the Auer-Mihalas (1972) model predictions. We now understand these to be limited by the incompleteness of the models (e.g., not taking low enough gravities for some of the hottest stars, incomplete physics (neglect of line-blanketing and wind backwarming), and uncertainties in the data (± 0.1 in log W)). On the other hand, this is the largest internally consistent body of data for O-type stars. Simon et al. (1983) compiled line-profile data on five early O-type stars, using a more complete set of models (and with some improvements in the non-LTE Auer-Mihalas calculations). Their data are considerably better than mine. These efforts are representative of T_{eff} derived from absorption lines in stars.

Another method of estimating T_{eff} in O-type stars comes from measuring the angular diameter and applying Equation (3-1) in the form:

$$\sigma\, T_{\rm eff}^4 = f/(\theta_d/2)^2 \quad , \qquad (3\text{-}2)$$

where θ_d is the measured angular diameter, and f is the total stellar flux at the Earth corrected for interstellar extinction. We have

$$f = \int_0^\infty f_\lambda d\lambda = \int_0^{\lambda_1} f_\lambda d\lambda + \int_{\lambda_1}^{\lambda_2} f_\lambda^{obs}\, 10^{0.4A_\lambda} d\lambda\, f \int_{\lambda_2}^\infty f_\lambda d\lambda \quad , \qquad (3\text{-}3)$$

where the RHS is divided into unobservable and observable portions (f_λ^{obs} between λ_1 and λ_2), respectively. The f_λ are not generally available below ~1200 Å (and cannot be obtained at all for $\lambda < 912$ Å) or above ~7000 Å. Models with predicted πF_λ (at the star) must be used to estimate the first and third terms of Equation (3-3).

Code et al. (1976) combined measured angular diameters θ_d for three O-type stars and a number of B stars to determine $T_{\rm eff}$. These are model-dependent to the extent that a

$$\int_0^{\lambda_1} \pi F_\lambda d\lambda$$

model integration needs to be made. The allowance for the long-wavelength contribution (beyond λ ~ 7000 Å) is small. The interstellar extinction correction A_λ, is appreciable in wavelength regions up to ~2000 Å. Interferometric measurements with present equipment can be made only for the handful of nearby stars, as discussed by Code et al.

Blackwell and Shallis (1977) pioneered a method for *estimating* θ_D by using measurements of f_λ and model predictions of F_λ. This method has been used extensively by Underhill (1982, and references therein) and has been commented on by Remie and Lamers (1982) and Tobin (1983). Write:

$$\theta_\lambda = 2\,(f_\lambda/\pi F_\lambda)^{1/2} \quad , \qquad (3\text{-}4)$$

where θ_λ is the angular diameter at a given λ. This can be estimated from the extinction-corrected f_λ and the model fluxes, πF_λ. One can write:

$$\sigma\, T_{\rm eff}^4\, \theta_D^2 = 4G \int_{\lambda_1}^{\lambda_2} f_\lambda^{obs}\, 10^{0.4A_\lambda}\, d\lambda \quad , \qquad (3\text{-}5)$$

where the G function is the separated explicit model-dependent portion:

$$G = 1 + \left(\int_0^{\lambda_1} \pi F_\lambda d\lambda + \int_{\lambda_2}^\infty \pi F_\lambda d\lambda\right) \Big/ \int_{\lambda_1}^{\lambda_2} \pi F_\lambda d\lambda \quad . \qquad (3\text{-}6)$$

Typically, λ_1 is 1200 Å and λ_2 is 7000 Å.

Thus, one measures f_λ^{obs} over an available wavelength range, adopts an initial model $T_{\rm eff}$, and uses these models to determine the G function. One then solves for a new $T_{\rm eff}$ by applying Equation (3-5), using an iterative procedure. The assumptions of this method are that f_λ is proportional to πF_λ at some convenient visible wavelength where $\theta_D = \theta_\lambda$ (this is probably all right for most O stars, but certainly not for Wolf-Rayet objects) and that G (model) represents real stars over all wavelengths. This is probably not a bad assumption for B stars for which the radiation below λ_1 is small in any case, but its applicability to hotter stars is dubious because the models are still quite inadequate physically (Abbott and Hummer, 1985). Nevertheless, I will tentatively adopt the Remie and Lamers (1982) values for the O and B stars measured by Underhill and her associates. (The reader may refer to Part III of this volume for Underhill's alternative numbers.)

Another related method for obtaining an effective temperature is to use those O stars in H II regions in which the Lyman continuum flux may be obtained from the nebulae to determine the total integrated flux. This method was pioneered by Morton (1969), and a modern and more complete discussion for eight O and 16 B stars is provided by Pottasch et al. (1979). This method avoids the complete dependence on models for the contribution to f_λ below λ_1, although it still depends on the interstellar extinction correction and potential pitfalls in nebular dust and density-bounded conditions. I will adopt the Pottasch et al. $T_{\rm eff}$ for their stars.

Finally, Abbott and Hummer (1985) have written a critique of all methods which use continuum estimates in the observable wavelength regions to determine T_{eff} for O stars. They point out that these methods are using only the Rayleigh-Jeans "tail" to determine the main features of the spectrum. They show that, in the case of ζ Pup, Code et al. (1976) found a T_{eff} that was 10000 K too low because of this fundamental uncertainty. The term,

$$\int_{\lambda_1}^{\lambda_2} \pi F_\lambda d\lambda$$

varies by only a factor 3 for models between 30000 and 50000 K; it depends strongly on gravity, and finally, backscattering must be included for a star with such a strong wind.

Nevertheless, keeping in mind all these potential problems, I list in Table 3-3 the O–Bl stars with T_{eff} determinations from Simon et al. (1983), Pottasch et al. (1979), Remie and Lamers (1982), Code et al. (1976), and Abbott and Hummer (1985). Five stars have been studied by more than one group and method. Aside from ζ Pup, the agreement is $\sim \pm 10$ percent. These values are plotted in Figure 3-2, which shows the individual points, with different symbols for "main-sequence" (type IV,V), giants (II,III), and supergiants (I,f) separately plotted. The dark solid line is that of Conti (1975) from the Bl to O6.5 subtypes, where the previous scale diverged to hotter T_{eff} (short dashed line). The extended dark solid line is drawn as a best current guess for the hot O stars. As can be seen, this line is still a reasonable current estimate of T_{eff} for main-sequence O stars. One could drop the T_{eff} relation by 10 percent in the middle portions to include the two O6.5 stars at $T_{eff} \sim 37000$ K if desired, although this would then put some "kinks" in the spectral-type T_{eff} relation. My best current estimate of T_{eff} for main-sequence O stars is the solid line of Figure 3-2, to ± 10 percent. Better models will ultimately help considerably, as would more systems studied by Pottasch's method.

Table 3-3
O–B1 Stars with T_{eff}, Bolometric Corrections from Individualized Measurements

		T_{eff}	-b.c.	Notes*
HD 93129A	O3 If	45000	4.20	1
HD 93250	O3 V	52500	4.71	1
HD 9128	O3 V	48000	4.35	1
HDE 303308	O3 V	45500	4.33	1
ζ Pup	O4 If	42000	4.43	1,5,6
9 Sgr	O4 V	44700	3.93	2
NGC 2244	(O4.5 V)	43600	3.84	2
HD 152723	O6.5 III	36800	3.35	2
HD 42088	O6.5 V	36800	3.35	2
HD 206267	O6.5 V	36200	3.30	2
+60 2522	O6.5 ef	33500	3.12	2
ξ Per	O7.5 III	32500	3.05	2
λ Ori	O8 III	31300	2.98	2
15 Mon	O8 III	31400	2.99	2
9 Sge	O8 If	34900	3.20	3
HD 151804	O8 If	33500	3.05	3
τ CMa	O9 I	33500	3.06	3
δ Ori A	O9.5 I	27600	2.67	3
ζ Ori A	O9.5 I	29900 25600	2.93 2.39	4, 3; 7
HD 188209	O9.5 I	27700	2.54	3
α Cam	O9.5 Ia	25600	2.45	3
AE Aur	O9.5 V	29500	2.82	2
ζ Oph	O9.5 V	31900	2.96	4
ϵ Ori	B0 Ia B0.5 Ia	26400 24800	2.44 2.46	3, 4; 7
μ Nor	B0 Ia	27800	2.57	4
15 Sgr	B0 Ia	32900	3.02	4
64 Cyg	B0 Ib	25200	2.32	3
τ Sco	B0 V	31800	3.00	3
κ Ori	B0.5 Ia	23100 26400	2.13 2.26	3, 4; 7
HD 150898	B0.5 Ia	24400	2.23	3
HD 167756	B0.5 Ia	25600	2.35	3
HD 64760	B0.5 Ib	24600	2.26	3
β Cru	B0.5 III	27600	2.75	4
λ Lep	B0.5 V	28200	2.75	3
δ Sco	B0.5 IV	31460	3.07	4
γ Ari	B1 Ib	20600	1.95	3
139 Tau	B1 Ib	21400	2.01	3
ζ Per	B1 Ib	19900	1.83	3
κ Cas	B1 Ia	21600	1.95	3
ρ Leo	B1 Iab	20300	1.84	3
β CMa	B1 II-III	24700 25200	2.33 2.38	3, 4; 8
15 CMa	B1 III	25100	2.39	3
β Cep	B1 III	26200	2.43	3
ϵ Cen	B1 III	25700	2.54	4
α Vir	B1 IV	23930	2.44	4

*Notes: 1 – Simon et al. (1983); 2 – Pottasch et al. (1979); 3 – Remie and Lamers (1982); 4 – Code et al. (1976); 5 – Abbott and Hummer (1985); 6 – Reference 3 finds T_{eff} = 35100, b.c. = -3.21; Reference 4 finds T_{eff} = 32500, b.c. = -3.18; all three values plotted; 7 – individually plotted; 8 – averages plotted.

Figure 3-2 shows that supergiants are 10 percent cooler at a given spectral type. This is also true for the O4f star, ζ Pup. This rule of thumb will probably also give a $T_{\rm eff}$ accurate to ± 10 percent. Giants may be intermediate, but individual points at spectral types Bl III and O8 III do not give confidence in this kind of consistency. The point for ζ Pup at 42000 K (Abbott and Hummer, 1985) is probably the best modeled in the diagram. For convenience, Table 3-4 lists the $T_{\rm eff}$ of O–Bl stars, taken from Figure 3-2. These are not the last word, and they should be used with caution. The bolometric corrections (b.c.) will be discussed in the next section; the M_v are listed for convenience. These have been adapted and interpolated from Table 3-1 by *assuming* that M_v should change relatively smoothly with spectral type. In some cases, the M_v values are different from the entries in Table 3-1 at a given spectral type, again indicating uncertainties in the calibration of some $0^m.3$ in the mean values.

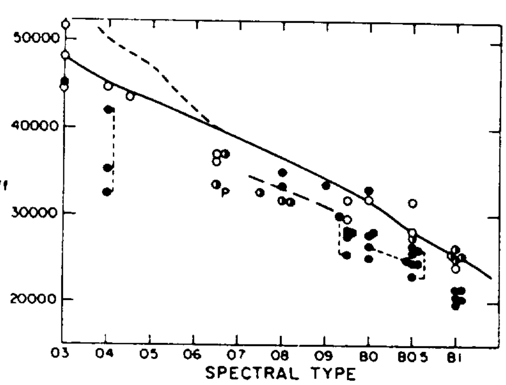


Figure 3-2. T_{eff}/spectral-type relation for OB stars. The individual points (open circles denote Type V; half-filled circles denote type III; filled circles denote type I or f) are taken from several references (see text). Symbols connected by the dashed lines refer to different T_{eff} for the same star. The long solid line is the best current estimate a mean T_{eff} for main-sequence stars. It is essentially identical to that of Conti (1973b) except for earlier than O6.5, where the dashed line shows his prior hotter scale for earlier stars.

All in all, then, the effective temperature scale for O stars, as estimated in Table 3-4, should be accurate to ± 10 percent since it agrees with line-ratio analyses, actual angular diameter measurements, the Blackwell and Shallis (1977) model-atmosphere approach, and determinations from H II region Lyman continuum flux measurements. For stars with substantial winds, such as ζ Pup, the "continuum" methods will not work well because wind backscattering plays a role, but Abbott and Hummer (1985) have modeled this star to have $T_{\rm eff}$ at 42000 K.

For the hottest O stars, in which most of the flux is emitted below the Lyman limit, the first variable term in the parentheses in Equation (3-6) will dominate, and the "continuum" methods will be very uncertain, except in those cases in which the f_λ can be estimated from a surrounding H II region. Abbott and Hummer (1985) have also called attention to this fact. Furthermore, all of the "continuum" methods depend on accurately estimating the UV extinction at short wavelengths. The usual procedure is to adopt a reddening "law" (e.g., Savage and Mathis, 1979) and determine the extinction at short wavelengths from $E(B-V)$ and/or "nulling" the λ2230 interstellar feature. Although this method may work in some cases of low extinction, it may be seriously in error in cases in which the UV extinction is anomalous. This seems to be a common phenomenon in our Galaxy, (e.g. Massa et al., 1983; Garmany et al., 1984; Massa and Savage, 1985). The reddening law *in the visible* may well be more or less the same from point to point in the Galaxy, but it seems clear that the UV extinction, which depends on a different-sized interstellar dust particle than that producing visible extinction, may behave differently. The adoption of a universal UV extinction was undoubtedly why Underhill (1982), using the Blackwell-Shallis procedure, obtained widely divergent $T_{\rm eff}$ for early O-type stars of the same spectral subclasses. All these stars are relatively heavily reddened: $E(B-V)$ is at least $0^m.35$ and is often considerably larger. Part III of this volume

Table 3-4
Estimated[a] Parameters of O–B1-Type Stars

Spectral Type	Luminosity class: V			III[b]		I/f		
	T_{eff}[c] (K)	–b.c.[d] (mag)	$-M_v$[e] (mag)	T_{eff}[c] (K)	$-M_v$[e] (mag)	T_{eff}[c] (K)	–b.c.[d] (mag)	$-M_v$[e] (mag)
O3	48000	4.4	5.6	45000	6.2	43000	4.2	6.2
O4	45000	4.0	5.4	43000	6.1	41000	3.8	6.4
O5	43000	3.7	5.2	41000	6.0	39000	3.5	6.5
O5.5	42000	3.5	5.15	40000	5.9	37500	3.3	6.5
O6	41000	3.4	5.1	39000	5.8	36500	3.2	6.5
O6.5	40000	3.3	5.0	38000	5.7	35500	3.2	6.5
O7	39000	3.2	4.9	37000	5.6	34500	3.2	6.5
O7.5	37500	3.2	4.8	35000	5.6	33000	3.1	6.5
O8	36500	3.1	4.6	34000	5.5	32000	3.0	6.5
O8.5	35000	3.0	4.5	33000	5.4	31000	2.8	6.3
O9	34000	3.0	4.4	32000	5.3	29500	2.7	6.1
B0	31800	2.9	3.9[f]	29500	5.0[f]	27000	2.4	6.0[g]
B0.5	28200	2.8	3.7[f]	26500	4.8[f]	25000	2.2	––
B1	25500	2.4	3.5[f]	23500	4.0[f]	21500	2.0	6.2[g]

[a]Uncertainties are still ± 10 percent.

[b]Interpolated between V and I/f entries.

[c]From Figure 3-1.

[d]From Figure 3-2.

[e]Smoothed from Table 3-1.

[f]From Lesh (1968).

[g]From Balona and Crampton (1974).

discusses this work and Underhill's rather different conclusions.

B. Wolf-Rayet Stars

I have already noted that the T_{eff} for Wolf-Rayet stars are highly uncertain. An absorption-line spectrum is not generally seen, but only the emission (line and continuum) from the wind is observed. Thus, the choice of appropriate models is critical. Physically complete theory is well in the future; a minimum requirement is for nonstatic models with extended geometry and statistical equilibrium calculations such as those of Hillier (1983). From the He II emission-line spectrum of the WN5 star, HD 50896, he found that a model with $T_{eff} \sim$ 60000 K fit the observations reasonably well. Even so, he notes that the fitting technique is not very sensitive to the temperature, which could well be higher.

Somewhat higher effective temperatures are inferred for the WN5 star, HD 193576 (also known as V444 Cyg), an eclipsing binary with well-determined orbital inclination. From an analysis of the light curve of the system, Cherepashchuk et al. (1984) found an electron temperature of 90000 K at the "core" (of 2.9 $R_{\odot}$) of V444 Cyg, implying a similarly high T_{eff}. In a recent paper, Pauldrach et al. (1985) report an improved code for calculating radiatively driven wind models. Their additions to previous work include two critical features: (1) an improved force multiplier term (Chapter 4), and (2) a finite angular sized radiative stellar core for calculating the line force. These authors could then calculate a moving wind model for V444 Cyg. They were able to duplicate the inferred density, velocity, and radius relationships derived by Cherepashchuk et al. for the WN5 star. Furthermore, they find consistency of the high T_{eff} of 90000 K. Thus we, now have a radiatively driven wind, and an empirical (eclipse data) model, in good agreement.

Considerable previous work on T_{eff} for Wolf-Rayet stars had inferred temperatures closer to 30000 K (e.g., see Parts I and III of this volume). Why the discrepancies? One point is that much of that work has used methods which involve the shape of the continuum over wavelengths from the observable FUV to visible regions. Application of the principles referred to previously, particularly Equations (3-5) and (3-6), leads to estimates of T_{eff}. One immediate concern is the use of the plane parallel LTE models of Kurucz (1979) for Wolf-Rayet stars in using this procedure. Furthermore, the UV extinction in our Galaxy is not a universal function, and substantial point-to-point variations are found (e.g., Massa and Savage, 1985). As Garmany et al. (1984) have noted, substantial errors in inferred Wolf-Rayet continua in the FUV can result when a uniform extinction law is adopted.

Morton (1970) has investigated the "Zanstra" temperatures of five single WN stars which are embedded in distinct H II regions. He estimated the Lyman continuum flux by using measurements of the free-free radio emission of the surrounding nebulae. He then obtained a value for the ratio of Lyman continuum flux to visible flux from the stellar apparent magnitude. This ratio is compared to models which are then used to determine the T_{eff}. He finds values near 50000 K for two WN5 stars, 33000 K for one WN6 star, and 23000 K for one WN8 star. A higher T_{eff} for another WN6 star is disregarded because he believes that other exciting stars in RCW 104, surrounding the star, HD 147419, have not been accounted for. This method is straightforward, since a real value of the Lyman continuum flux is used; however, the models used (Morton, 1969) are LTE-blanketed ones with plane parallel geometry, with no accounting for the extensive wind and its backscattering. Therefore, I find the numbers for T_{eff} to be interesting but very unlikely to be correct because of this severe model-dependence. The eclipsing binary data are nearly model-independent and are to be preferred. The factor of nearly 2 discrepancy between the Zanstra inferred values near 50000 K and the 90000 K from the eclipsing binary data for WN5 stars indicates the depth of our ignorance concerning Wolf-Rayet temperature parameters, particularly the Lyman continuum predictions.

IV. BOLOMETRIC CORRECTIONS

A bolometric correction (b.c.) is used to correct an absolute visual magnitude, M_v, for the total energy (radiative and nonthermal) emitted at all other wavelengths. The resultant M_{bol} is then an estimate of the total luminosity in ergs s^{-1}. For very hot stars, such as those that we are dealing with here, the b.c. can be substantial, completely dominating the resultant luminosity, since our visual observations are out on the tail end of the emergent continuum. Observed b.c.'s of 3 to 5 magnitudes, or factors of 16 to 100, are to be expected in these stars.

A. O-Type Stars

The b.c. for O-type stars must be estimated from appropriate models because a substantial fraction of the emergent flux will come from the mostly unobservable FUV regions near the Lyman limit and beyond. For stars embedded in H II regions in which the Lyman continuum photons are bounded and in which the dust absorption can be neglected, the Balmer photons can be used to estimate the Lyman flux. This will help to lessen the uncertainties in the bolometric corrections.

To obtain a rough estimate of b.c. for O-type stars, I will use the same procedure as that of Section III; namely, combine the work of Simon et al. (1983), Code et al. (1976), Remie and Lamers (1982), and Pottasch et al. (1979). These authors provided b.c.'s for those stars for which they list the T_{eff}. Taking all these numbers at face value, I have plotted them versus spectral type in Figure 3-3. The agreement among the various authors is not bad, considering the divergent methods. The late-type supergiants, being cooler than the corresponding main-sequence stars, also have smaller bolometric corrections. I have drawn a line through the main-sequence points; it is a smooth function of spectral type. At type O6.5, I have arbitrarily adopted the Pottasch et al. listed values, admittedly somewhat inconsistently with the choice I made in Figure 3-2. The b.c. estimated here are probably accurate to $\pm 0^m.3$.

The late OB-type supergiants have b.c. smaller by $0^m.5$ than the corresponding main-sequence stars. There are no direct determinations of b.c. for earlier type O supergiants and Of stars except for ζ Pup. Its b.c. appears to be larger than an O4 main-sequence star, 9 Sgr, but this may represent the different methods of Simon et al. (1983) and Pottasch et al. (1979), respectively.

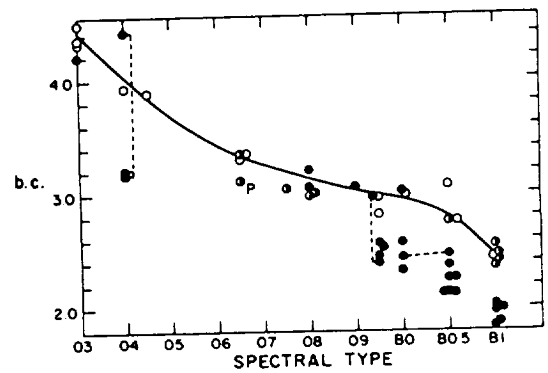


Figure 3-3. Bolometric corrections (b.c.)/spectral-type relation for OB stars. Symbols are as in Figure 3-2. The solid line indicates the best current estimate for the relationship for main-sequence stars.

I have used Figure 3-3 as an interpolation diagram to derive the values listed in Table 3-4. The b.c. numbers for intermediate luminosity cases can be interpolated between the entries for luminosity classes V and I, although Figure 3-3 indicates that the numbers may be closer to type V. The mean M_{bol} of the brightest and earliest O3f stars is $-10^m.4$ or $L \sim 10^6 L_\odot$.

B. Wolf-Rayet Stars

Because no well-established T_{eff} scale exists for Wolf-Rayet stars and the models are thus far mostly inadequate, it is not surprising that the b.c. are not well known. Estimated values may be found in Chapter 1 of this volume, but these are based on plane parallel models and a relatively cool temperature scale, compared to the concerns of Section III above. We might expect fairly substantial b.c. if T_{eff} are as high as 90000 K, as is indicated there for some stars.

Indirect evidence for substantial b.c. for Wolf-Rayet stars does exist, but from a different quarter. If, as appears likely, Wolf-Rayet stars are helium-burning descendants of relatively massive stars, one can appeal to evolution arguments and cluster/association membership of Wolf-Rayet and massive O-type stars. This exercise has been carried out by Humphreys et al. (1985), who discuss galactic Wolf-Rayet star membership in known associations. They determine the M_v of the most massive O stars, add the b.c., and determine the M_{bol}. Under the assumption that the Wolf-Rayet stars are evolved from stars at least as massive as these O types, they find a *minimum* b.c. by measuring the difference in magnitude between the M_{bol} of the O stars and the M_v of the Wolf-Rayet star(s). These b.c. are typically 4 to 6^m for the 24 or so Wolf-Rayet stars with well-established membership in associations. The sample does not cover all spectral subtypes; however, the later WN types, with brighter M_v (Table 3-2), have somewhat smaller minimum b.c. The implied b.c. for the other spectral subclasses are large enough to reinforce our view that the $T_{\rm eff}$ scale is considerably hotter than the 30000 K quoted in Chapter 1 of this volume.

The b.c. listed by Humphreys et al. (1985) are very rough estimates and at least one step removed from real Wolf-Rayet star models because they depend on O-type stellar atmospheres. These values of 4 to 6^m should be considered only as guides to true values. When Wolf-Rayet effective temperature scales are realized and appropriate stellar atmosphere models are constructed, we may have some real understanding of the total M_{bol} for these objects. Since I have not listed $T_{\rm eff}$ for Wolf-Rayet subtypes, it is also imprudent to list b.c. for the different classes. It seems likely that the b.c. are substantial and may depend on subtype, being larger for earlier excitation classes than for later ones. It is even possible that the M_v-spectral-type relations of Figure 3-1, which show *fainter* M_v at earlier types, might have less different M_{bol} at all subclasses.

V. RADII FROM ECLIPSING BINARIES

Studies of eclipsing binaries can give direct information on the two-component stars because photometry, particularly at different wavelengths, can give eclipse durations and thus estimates of relative radii. Photometric variability, in and out of eclipse, gives constraints on relative luminosities, and radial velocities lead to orbital solutions and relatively complete estimates of the relevant stellar parameters. These detailed techniques, although quite powerful, can be applied only to those fortuitous systems in which the Earth's line of sight is near the orbital plane. Complications can arise because of interactive phenomena involving the close companions and nonspherically symmetric winds with gaseous streams. Only a handful of O and Wolf-Rayet eclipsing binary systems have been studied in detail, both photometrically and spectroscopically, and I will briefly discuss them here insofar as the data affect information concerning stellar radii.

A. O-Type Stars

1. 29 UW CMa. Leung and Schneider (1978a) summarize previous work on this system and discuss a complete photometric solution with their new technique. They find the O7 If + OB stars to be in contact with radii of 20.4 and 17.8 $R_\odot$, respectively. The radial-velocity curves give a spurious eccentricity since the photometric solution implies a circular orbit. This is probably due to tidal distortions and reflection effects which are often found in close-contact systems. Drechsel (1978) also found R(Of) $\sim 20.4\ R_\odot$ from these photometric data. The UV spectrum of this system has been discussed by Drechsel et al. (1981) and references therein, on the basis of data from Copernicus and IUE satellites.

2. AO Cas. Schneider and Leung (1978) reanalyze previous photometry and list references to existing work. They find this to be a contact system with two O8.5 III stars whose radii are 13 and 14 $R_\odot$. The secondary is predicted to be slightly brighter, although its observed absorption lines are weaker than those of the primary.

3. V729 Cyg = BD+40 4220. Leung and Schneider (1978b) summarize and discuss the photometry and spectroscopy of this system, a member of the Cyg OB2 association. They find radii of 33 and 16 $R_\odot$ for the primary and secondary, both O7f stars. The secondary is predicted to be fainter, yet the absorption lines in each component are similar; hence, as in AO Cas, a discrepancy in the photometry/line strengths is seen, although in the opposite sense.

The large radii of these systems seem to be well established, even with the difficulties mentioned above. The luminosities or implied T_{eff} may not be well determined. In any case, the radii of O stars well removed from the zero-age main sequence (ZAMS) are roughly consistent with their types. There is no evidence for radii that are greatly dependent on wavelengths, unlike the situation for Wolf-Rayet stars, as we shall presently see. However, Eaton (1978) found that the O7f component of UW CMa was slightly larger in FUV wavelengths. An interesting recent study of the variable Hα profile of this system has been provided by Vreux (1985).

B. Wolf-Rayet Stars

Cherepashchuk (1982) gives a nice review of eclipsing binary work on the galactic Wolf-Rayet systems, V444 Cyg, CX Cep, CQ Cep, and CV Ser. He lists core radii of 2.9, 4.5, 11, and 3-4 $R_\odot$ for the WN5, WN6, WN7, and WC8 spectral types, respectively. These core radii imply correspondingly high effective temperatures.

Subsequently, Leung et al. (1983) rediscussed the parameters of CQ Cep and obtained radii of 8 to 13 $R_\odot$ for different assumptions of the mass ratio, in reasonable agreement with previous work. A detailed analysis of this binary system was recently published by Stickland et al. (1984), who conclude that this is a contact system with the inferred O star immersed in the wind of the WN7 primary. Because the system is a likely member of the Cep OB1 association with a known distance, they have a good constraint on the system parameters. Using the flux over all measurable wavelengths and rough continuum fits to blackbody predictions, they derive a T_{eff} of 30000 K and a core radius of 16.7 $R_\odot$ for the WN7 star and a radius of 14.4 $R_\odot$ for its companion. The total radii sum to considerably more than the mutual separation of 20 $R_\odot$, which is derived from the orbital parameters and an assumption about the mass ratio. (The secondary is not observed.) Stickland et al. point out that the discrepancy can be resolved only by making either or both components considerably smaller and therefore hotter. They argue that the flux-fitting procedure is probably in error and that an appreciably hotter and smaller radius for the WN7 star will then be consistent with all the eclipse and spectroscopic data. They suggest a best fit of ~ 11 $R_\odot$ for both components, with the Wolf-Rayet star being hotter and thus brighter. This value is similar to that listed by Cherepashchuk (1982).

Cherepashchuk et al. (1984) have also recently derived the structure of the expanding atmosphere and the physical characteristics of the WN5 component of V444 Cyg, using UV photometry from OAO-2 (Orbiting Astronomical Obseratory) and light curves from 2460 Å to 3.5 μm. They determine a core radius of 2.9 $R_\odot$ and a corresponding temperature at this level of 90000 K. They find that the eclipse curves are very dependent on wavelength; the shape of the primary minimum remains fixed while its width increases with increasing wavelength (in primary minimum, the WN5 component is eclipsed by the O star), confirming a result found many years ago by Kuhi (1968).

An innovative wavelength-dependent light-curve analysis (Cherepashchuk, 1973) is used to constrain the eclipse solutions. A relatively cool color temperature, 20000 K, is found from the UV continua, once interstellar extinction is estimated; an 8000 K infrared color temperature is deduced. These values are typical of Wolf-Rayet continua and should not be confused with core temperatures, or T_{eff}, which are clearly considerably higher. The continuum-fitting procedure for Wolf-Rayet stars consists of sampling different levels in the wind, not the intrinsic parameters of the star.

The eclipsing binary work done so far points strongly to relatively small core radii and relatively high effective temperatures for Wolf-Rayet stars. These estimates are mainly model-independent and are much more preferred than other work which appeals to continuum measurements, uncertain corrections for radiation below this Lyman limit, assumptions concerning uniform FUV extinction, etc.

VI. COMPOSITION

A. O-Type Stars

The chemical composition of stars can be inferred from their spectra, using a detailed analysis of the line strengths with prior knowledge of the appropriate atomic physics parameters, including the transition probabilities. The background must include a model of the appropriate stellar atmosphere with knowledge of the effective temperature and surface gravities.

Let me assert at the outset that, for O-type stars, the models are probably adequate for obtaining abundances to better than a factor of 2 *if* the appropriate atomic physics parameters are known. These are well established for hydrogen and helium transitions, but not so well in hand for all the CNO and silicon ions which make up the bulk of the other absorption lines found in O stars. Although there have been no "full-blown" detailed analyses of the line spectra of many O stars, bits and pieces of the problem have been attacked. Basically, the spectra of O-type stars appear to indicate mainly normal (solar) composition, except for the ON and OC subtypes. In addition, Brown et al. (1986) have found an apparent deficiency of nitrogen in the O stars of NGC 6611, compared to those in NGC 6231 (Keenan et al., 1984).

In what follows, it should be kept in mind that the atmospheric models are still only approximations for real stars, but are probably applicable for composition studies. Conti (1973a) found that models with normal hydrogen/helium ratios (10:1) gave a reasonable fit to the equivalent width data when appropriate effective temperatures and gravities were derived. Other choices of H/He ratios would lead to rather different T_{eff} and g_{eff}, giving inconsistent inferred masses. Because all O-stars fit a normal H/He ratio, it is presumed that they do not differ among themselves.

In a more detailed study of some early O-type stars, Conti and Frost (1977) found normal H/He ratios for all stars except ζ Pup, which had strong 5-n series He II lines. They speculated that helium might be somewhat enhanced in the latter object. In an improved spectroscopic study involving a better non-LTE code, Kudritzki et al. (1983) also found an enhanced helium composition for ζ Pup. Analysis of all the optical lines indicates that this may in fact be the final answer on this issue, but I must confess that, at the level of significance found (a 30-percent enhancement), I am nervous about the adequacy of the models. The new Abbott and Hummer (1985) model for ζ Pup introduces a different structure to the atmosphere. It is therefore possible that the deduced enhanced helium abundance is an artifact of model inadequacies. At present, there is no definitive conclusion on this issue.

Takada (1977) published a detailed spectroscopic study of two O supergiants. Using planar non-LTE models, he found that normal H/He compositions fit the data. He found that the ions from other common elements also appeared to be consistent in strength with other early-type stars, except for neon, which was enhanced.

Sakhibullin et al. (1982) and Sakhibullin and van der Hucht (1984) have carried out a detailed analysis of carbon ion-line transitions in a number of O-type stars. They make use of Copernicus FUV data and visible data in the literature. For most stars, the abundance appears to be normal.

I have already remarked on the ON stars with strong nitrogen lines and the OC stars with strong carbon lines. They also appear to represent anomalous abundances, although the quantitative values are not well established. In particular, it is not certain whether the ON stars are a distinct nitrogen-enriched class of objects or only the most extreme examples of a continuous range of nitrogen abundance from normal. Also, it has been suggested that the OC supergaints might have anomalous atmospheric conditions, being very extended, and thus exhibiting the strong carbon lines (López and Walsh, 1983).

Some quantitative work on a single ON star is given by Lester (1973), who found nitrogen overabundant and carbon deficient in HD 201345 relative to 10 Lac. In a modern analysis of a few stars, Schönberner et al. (1984) found nitrogen overabundances and, furthermore, helium also enhanced. Kudritzki and Hummer (1986) point out that the values observed are similar to "incomplete" CNO-cycle equilibrium. The UV spectral morphology of ON and OC stars is illustrated by Walborn and Panek (1985). It seems clear that abundance anomalies are involved, but their origin is not yet certain. I would guess that mixing inside the star has played a role. More quantitative work is needed, particularly with respect to the issue of discrete classes versus continuous properties.

B. Wolf-Rayet Stars

Here we are dealing with a spectroscopic situation in which only emission lines from a wind are observed; the material must be representative of the underlying star, yet the models used thus far are clearly inadequate. On the other hand, the presence or absence of various ions in the spectra are so obviously different from normal that chemical anomalies appear to be the only reasonable conclusion, particularly since they mirror processes occurring in stellar interiors (Gamov, 1943).

Let us first address the hydrogen and helium abundances. In a careful visible-wavelength survey of 60 WN stars, Conti et al. (1983b) found that roughly one-third had evidence of hydrogen in that the hydrogen Balmer lines were stronger than the alternating Pickering (He II) lines. It was mainly the later WN-type stars that showed this phenomenon. All spectral subclasses of vastly different excitation had at least one case of a Balmer/Pickering decrement, indicating the presence of hydrogen and thus arguing strongly that the differences were not primarily due to local conditions, but must be real. How else could we explain why stars of the same subtype have different Balmer/Pickering ratios? This qualitative result would appear to be model-independent.

To assign actual numbers, one must adopt a model and parameters, the most important facet being an ionization temperature. Conti et al. (1983b) found a range of H/He ratios in the WN stars with *detected* hydrogen from 5 down to 0.2. They assigned upper limits of 0.1 to this ratio for the two-thirds of the WN stars without any appreciable enhancement of the Balmer lines with respect to the Pickering series. The upper limit value could be zero, but because the uncertainties in the actual numbers are probably a factor 2, ratios up to 0.2 are not excluded. It is clear that, for most WN stars, hydrogen is deficient because it is observed weakly or not at all. This conclusion was first pointed out by Smith (1973).

In the WC stars, the question of hydrogen is a little more clouded because C IV lines (and C III in some cases) overlap the Balmer lines. Nugis (1975) reported the apparent absence of hydrogen in several WC stars. Torres et al. (1986), in their survey of 80 WC stars, found no evidence of hydrogen from the visible spectrum. Willis (1982b) has carefully examined the FUV spectrum of the bright WC8 star, HD 192103, and found no evidence of Lyα. This is probably the most detailed study of any WC star and could be taken as being representative of all WC objects. Although the case is not as clearcut as that of WN stars, a reasonable conclusion is that hydrogen is not present in WC stars. This is consistent with evolutionary considerations (Chapter 5), but that is getting ahead of the story.

A nice review of the chemical composition of Wolf-Rayet stars is given by Willis (1982a). In WN stars, the nitrogen appears to be enhanced by factors of 10 to 100 (Nugis, 1975; Willis and Wilson, 1978; Smith and Willis, 1982; Conti et al., 1983b) and the carbon appears to be decreased by similar factors. These analyses have attempted to be model-independent insofar as comparing ions of similar excitation and ionization, but of course, they are not yet in final form. For this reason, I have chosen not to list values here. Needless to say, the great strength of the nitrogen ions and near absence of carbon ions (except C IV) in WN stars would seem to have its origin in abundance anomalies caused by the CNO cycle of core-hydrogen burning. Other ad hoc suggestions to account for the spectroscopic disparities appear in Part III of this volume.

In WC stars, the analyses (Nugis, 1975; Willis and Wilson, 1978; Smith and Willis, 1982; Torres et al., 1986) indicate that carbon and oxygen are overabundant, whereas nitrogen is absent (Nussbaumer and Schmutz, 1986; Willis et al., 1986). *No* ions of nitrogen are found in the UV, visible, or IR spectra of *any* WC-type stars. It is difficult for me to imagine a physical situation in which the nitrogen abundance is "normal," yet *no* lines of *any ions* of the element are seen in any spectral region. The actual numbers of carbon, as compared to helium, indicate a substantial enhancement of the element so that the C/He ratio is ≈ 0.1, and carbon is then the second most abundant element. Chapter 5 discussed how this is related in a natural way to the evolution of massive stars as they burn helium. Torres (1985) has found that the C/He ratio appears to be similar in WC stars of all subtypes.

Finally, we have implied that most WN stars and all WC stars are more or less devoid of hydrogen and should therefore be helium-burning objects. Thus, the two immediate consequences are: (1) they should be overluminous for their mass, and (2) they should have relatively high effective temperatures (to the left of the ZAMS). With respect to the former issue, it is well known (e.g., Paczynski, 1973) that several Wolf-Rayet stars in binary systems have masses of $\sim 10\ M_\odot$ but luminosities substantially in excess of those predicted for hydrogen-burning stars of this mass. With respect to the latter issue, from a detailed eclipsing binary analysis of V444 Cyg (Cherepashchuk et al., 1984), we now find (Section III) that the T_{eff} indeed seem to be hot. The final word on a T_{eff} scale for Wolf-Rayet stars is not yet in hand, but the most direct and model-independent evidence is for stars on or near the helium-burning sequence. Overall, then, the composition of Wolf-Rayet stars appears to be hydrogen-poor and helium-rich, with the WN stars enhanced in nitrogen and the WC stars enhanced in carbon. These will follow as normal consequences of evolution (Chapter 5).

Evolutionary models of massive helium-burning stars, with mass loss included, have been constructed, for example, by Maeder (1983) and Prantzos, et al. (1986). These authors find that, typically in WN stars, helium and nitrogen are enhanced, as is observed, and carbon and oxygen become prevalent in WC stars as core material reaches the surface. Qualitatively, the agreement is very impressive; quantitatively, there appear to be discrepancies of a factor of a few due to inadequacies in the data and in the modeling. The overall trends, however, are completely consistent between observation and theory.

VII. MASSES

A. O-Type Stars

Stellar masses can be determined directly only for binary systems in which both components are detected and the orbital parameters are established. This can be done for visible binaries and for double-lined spectroscopic systems which eclipse. Unfortunately, because no visible binary orbits are well determined for O-type systems, we must be content to deal with eclipsing double-lined spectroscopic binaries. The values for a number of systems have been discussed in Chapter 1 (Table 1-25), and I do not need to repeat them here. For double-lined systems which do not eclipse, only lower limits to the masses may be obtained due to the uncertainty in the orbital inclinations. However, even data of this type are useful in some circumstances.

Indirect inferences on stellar masses can come from placement of the stars in a theoretical HR diagram whence the comparison to the evolution tracks gives estimates of the masses. Conti and Burnichon (1975) went through this exercise for O stars, although the tracks used then were conservative (i.e., they included no mass-loss parameterizations). Since I will return to these themes later in Chapter 5, it is not necessary to include much more discussion here. It is helpful to point out that the agreement between the masses determined from relatively unevolved O-type binary systems and those inferred from the placement of stars on the theoretical evolution tracks is not bad. For some highly evolved systems (e.g., supergiant companions of X-ray sources), the stars seem to be undermassive for their luminosity (e.g., Conti, 1978). The most massive known system, HD 47129, has masses which are large (near 60 M according to Hutchings and Cowley, 1976) but which are consistent with their spectral-type placement in the HR diagram.

Detailed study of spectroscopic binary systems requires dedicated and lengthy observing programs, along with careful analysis. It is critical to obtain more data for more systems. Of particular importance will be the analysis of binaries found in the Magellanic Clouds, and preliminary work by Niemela (1986) on three systems has recently been reported.

B. Wolf-Rayet Stars

The spectroscopic analysis of these systems is often complicated by the fact that one is measuring emission lines in the Wolf-Rayet component which are formed in the stellar wind, along with absorption lines in the O-type companion (if it is seen) which are formed in a more static atmosphere. Frequently, the velocity amplitudes for various Wolf-Rayet emission lines differ appreciably; in most cases, those of highest excitation give the smallest values and are presumed to be representative of the deepest layers. The systemic velocities of the emission lines often differ from those of the absorption lines of the companion; the latter are believed to best represent the binary motion. Some Wolf-Rayet binary systems contain components in close contact with one another, and the interactions between the individual stellar winds lead to very complicated and phase-dependent variability. The interpretation of the radial-velocity curves and the inferred masses and mass ratios is therefore difficult.

In nearly all Wolf-Rayet systems, the Wolf-Rayet component is the less massive, consistent with its expected more advanced evolutionary state compared to its companion. A thorough discussion of the values of masses and mass ratios in Wolf-Rayet binary systems is given by Massey (1981), and the values are listed in Table 1-28 of Chapter 1. A few additional systems not yet studied in 1981 are given there. Progress in acquiring more data is very slow, but will be critically important in our understanding of the evolutionary status of Wolf-Rayet stars.

VIII. GROUND-BASED OBSERVATIONS OF INTRINSIC VARIATIONS IN O, Of, AND WOLF-RAYET STARS*

A. Introduction

Stars cannot be subjected to laboratory experiments. However, if a star is variable, the observer may passively participate in such

*Based partly on observations obtained by Baade at the European Southern Observatory, La Silla, Chile.

stellar self-experiments and, in parameter space, look at the star from varying points of view. Observations of variations should therefore provide deeper insights than do single snapshots. For Wolf-Rayet stars, particularly, this appears as an attractive possibility because, in most of them, only the rapidly expanding wind is visible. A necessary condition, however, is that the character of the variability is well known. Unfortunately, it cannot be claimed that this stage has been reached for any of the stars considered in this volume, but the recognition of some regular and semiregular patterns in their variations clearly marks a first step.

Instability, as an extreme case of variability, is probably the reason that stars with masses beyond a certain threshold do not exist (cf. the recent review by Appenzeller, 1986). Does variability (e.g., via altered or additional mass-loss processes) also affect the evolution of stars below that threshold? Again, we are far from being able to give an answer with any certainty, but the question will form part of the background to this section.

For several reasons, the intrinsic variability of O and Wolf-Rayet stars has not thus far been a very prominent topic. Apart from compiling the observational evidence of intrinsic variations of O and Wolf-Rayet stars, this section will therefore make a first crude attempt to categorize them, and attention will be paid to possible connections between different variabilities. Not included are variations which are most probably the result of a star's having a companion. The restriction to ground-based observations (except for flux measurements) excludes the variability of the high-velocity regions of stellar winds. This topic is covered by Hendrichs' chapter in this volume, and I will only try to provide some cross references where possible. Finally, I shall rather strictly observe the limits in spectral type even though the distinction between O and B stars is quite arbitrary, particularly with regard to variabilities associated with high luminosity and fast rotation.

Summaries of the pulsations and related variabilities in Wolf-Rayet stars and OB supergiants have been given by Maeder (1986), in B stars with and without emission lines by Percy (1986) and Baade (1986), respectively, and most recently, in OB stars in general by Smith (1986b). The workshop on *Instabilities in Luminous Early-Type Stars* (de Loore and Lamers, 1987) has treated its subject on a broad observational and theoretical basis, including a review by Vreux (1987) of the variability of Wolf-Rayet stars, whereas the earlier workshop on *The Connection Between Nonvalid Pulsations and Mass Loss* (Abbott et al., 1986) was devoted to a more specific problem.

B. O Stars

1. Variations at the Photospheric Level

a. Flux and Flux Distribution. For main-sequence O stars, hardly anything is known about the occurrence of photometric variations. This and the failure of searches for a blue extension of the β Cephei instability strip (B0.5 to B2) should mean that the amplitudes, if any, do not generally exceed a few hundredths of a magnitude. Surveys are strongly biased to supergiants, where variations are more readily found. From the comparison of their survey of OBA supergiants with the one by Klare and Neckel (1977), Schild et al. (1983) found significant magnitude differences for at least 20 percent of the stars in both surveys. Using the large homogeneous data base accumulated in the Geneva photometric system (Rufener, 1980), Maeder (1980) confirmed earlier results that, with mean amplitudes around $0^{m}05$, the level of variability in O supergiants is about the same as in B and A supergiants (with a possible peak near early B types). Because the amplitudes increase with luminosity, as is also the case for most of the rest of the Hertzsprung-Russell (HR) diagram, he also found that they correlated with the mass-loss rates. But no study has yet been undertaken as to whether this correlation also means a causal connection of the two phenomena or whether it is merely coincidental. (However, note the work by Noels and Gabriel (1984) and Maeder (1985), who find that high mass-loss rates are a necessary condition for maintaining radial pulsations in Wolf-Rayet stars.)

With a minimum of only three measurements as a selection criterion, Maeder's data cannot be used to estimate the fraction of variable supergiants. However, more detailed observations of individual stars show that, at the 0^m01 level, constancy over periods of days appears to be the exception rather than the rule. Such studies of several stars each have been made by van Genderen (1985, 1986), van Genderen et al. (1985a, 1985b), and Percy and Welch (1983). Their most important result is that the variations are semiregular and probably form an extension of the general quasi-period/color-luminosity relation derived by Burki (1978) for supergiants with spectral types between B1 and G0. Van Genderen and coworkers furthermore suspect that, in some cases, the variability consists of a quite regular, perhaps periodic, component, plus additional irregular fluctuations. Longer series of observations appear to be necessary for substantiating this suggestion.

O stars radiate most of their energy in the FUV region. It is therefore important that, from observations with the low-resolution spectrophotometer on board Voyager 2, Polidan et al. (1985) report flux variations below 1100 Å for a large fraction of their sample of OB stars. However, many stars were also perfectly constant within the observational limits, as confirmed by the study of a smaller sample with the same instrument by Polidan and Stalio (1985). From IUE low-resolution spectroscopy of four OB supergiants, Underhill (1984) found amplitudes of about 0^m1 also at the longer UV wavelengths.

Very little is known about color variations. From their comparison of two surveys, Schild et al. (1983) believe that the differences are not significant. Simultaneous five-channel photometry in the Walraven system permitted van Genderen (1986) to detect color variations at a lower level in a number of stars. Although they are consistent with temperature variations, in no case could global temperature changes alone account for the brightness variations. The problems encountered by van Genderen in deriving a consistent solution are reminiscent of the results obtained from the first applications of the Wesselink method to β Cephei stars (cf. Lesh and Aizenman, 1978) and could mean that the temperature variations are not uniform and/or that temperature and radius vary in antiphase.

The quasi-periods of early-type supergiants are often of the order of a few days and are therefore several times longer than the periods of the fundamental radial pulsation mode. This has led to the suggestion (Maeder, 1980b; Lovy et al., 1984) to identify these variations with nonradial pulsation in g-modes. Such modes could become unstable as a result of the large outer convection zones induced by the strong radiation pressure in very luminous stars. An important property of g-modes in this connection is that, in the presence of only convection, no real periods are associated with them. Only with an additional restoring force (such as rotation (Ledoux, 1969)) can truly periodic motions develop.

b. Radial Velocities, Macroturbulence, etc. Systematic searches for radial-velocity variations have been made in order to assess the binary frequency of O stars (Garmany et al., 1980, and references therein). An important byproduct was the detection of variable atmospheric velocity fields in many of the stars studied (Bohannan and Garmany, 1978; Garmany et al., 1980). In consequence of their original aim, the time resolution of these observations was generally insufficient to identify regular variations on a dynamical time scale below 1 day. The velocity ranges in several stars reached 30 km/s. Owing to projection effects, they would even be only a lower limit to the true velocity amplitudes if the variations were due to a uniform expansion. If all radial velocities pertain to the same mass layers, such large radial-velocity amplitudes are in conflict with the general absence of large brightness and color variations (Section VIII.B.a), unless the time scales are very short. It is therefore likely that the observed variations are due to inhomogeneous velocity fields (such as macroturbulence or nonradial pulsation), in which case the measured amplitudes would reflect mainly the star's rotation. Depending on how the measurements are made, this would also considerably increase the threshold (Garmany et al. used 45 km/s) beyond which velocity variations of unknown period are most reasonably explained by orbital motions.

High-resolution low-noise spectroscopy (Ebbets, 1982; Baade, 1983, 1984b; Baade and Ferlet, in preparation) of selected absorption lines in several bright supergiants with low and intermediate *V sin i* (the distinction from broad-lined supergiants may be important (cf. below)) detected variable profiles in nearly all stars observed. It seems likely that they correspond to the radial-velocity variations measured in photographic spectra (see above). The line profiles often display persistent as well as variable asymmetries. Eventually, it may become possible to identify some characteristic patterns of the variability, but in contrast to broad-lined supergiants (cf. Section VIII.B.1.c(1)), clear evidence of nonradial pulsations governing the line-profile variations has not yet been found. Only certain very subtle structures at the bottom of the profiles of some stars (see Figure 3-4) are now candidates of being due to nonradial pulsations. In any case, the pulsation amplitudes must be small.

Conti and Ebbets (1977) have studied the distribution of rotational velocities, *V sin i*, of 205 O-type stars. With this large a sample, the long known strong underabundance of narrow-lined O stars became statistically tractable. In particular, the distribution function for evolved stars deviates noticeably from model calculations at the low-velocity end. Therefore, another line-broadening mechanism probably exists in addition to rotation, and Conti and Ebbets follow Slettebak's (1956) suggestion of macroturbulence as its most likely identification. Both analyses agree that macroturbulence

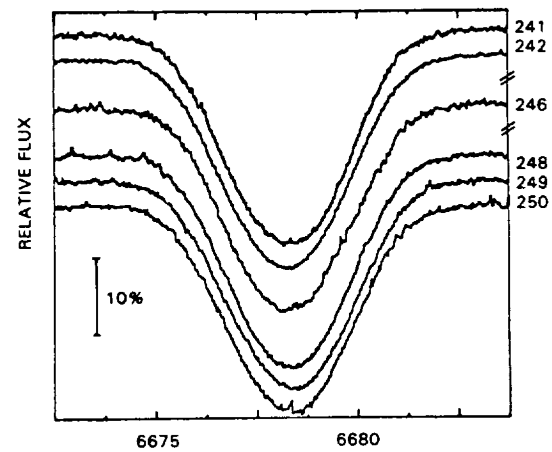


Figure 3-4. An example of absorption-line/profile variability in OB supergiants. Shown is the He I λ6678 line in θ Ara (B2 Ib); observing dates are labeled on the right-hand side as JD-2,446,000. The smooth symmetry-changing microvariations of the line core come relatively close to the expected effects of single-mode nonradial pulsation of very low amplitude. Other OB supergiants often show much larger variations of the entire line profile which are clearly not due to single-mode pulsation. (See also Figure 3-6b.)

increases toward earlier spectral types and more evolved stars. Slettebak had furthermore concluded that, in supergiants later than B1, line-broadening is by rotation only.

c. *Nonradial Pulsations*

(1) The Observational Evidence. Although nonradial pulsations (NRP's) are theoretically expected in highly evolved stars (Section VIII.B.1.a) but are observationally not definitively confirmed in narrow- and intermediate-lined supergiants (Section VIII.B.1.b), the situation is quite the opposite near and on the main sequence. The first spectroscopic discovery of NRP's in any O star was by Smith (1978, 1981) in the nonsupergiants, 10 Lac (O8 III) and ν Ori (O9.5 V), but so far, theoretical model calculations have not uncovered a convincing driving mechanism (Osaki, 1986a, and references therein).

The initial impression that these stars form the hot end of a separate new group of variable OB stars will probably have to be replaced by a picture in which the majority of normal OB stars earlier than about B8 pulsate nonradially at least occasionally (cf. Baade, 1986). Until now, however, only very few O stars have been searched for NRP's so that the frequency of occurrence of NRP's in the HR diagram is basically unknown. It cannot be generally very low because otherwise the proportion of nonradially pulsating O stars would not be so large (at least 25 percent) as it is in the current sample.

Smith's (1978, 1981) observations of the narrow-lined stars, 10 Lac and ν Ori, showed small line-profile variations, indicating velocity amplitudes of a few km/s. Single-mode line-fitting suggested time scales between a few hours and 1 day. In ν Ori on one occasion, a variation by a factor of 2 was found, and the behavior of 10 Lac was not quite regular. Later experience gained from stars with larger *V sin i* and variable amplitudes of two or more modes would not rule out a more complex mode spectrum but with constant periods (Smith, 1986a).

That broad lines can convey more information about the nonradial pulsations of a star was demonstrated by Vogt and Penrod (1983), who used the imaging effect of the rotational Doppler line-broadening to diagnose, from moving multiple bumps in the line profiles, an $\ell = 8$, $m = -8$ mode in ζ Oph (O9.5 Ve). At the same time, their observations (spanning 2 years) are a good illustration of how variable the appearance of the phenomenon can be in a single star. In particular, the amplitude of the pulsation can vary to render the bumps nearly invisible. But it is also evident that the pulsational waves traveling about the star at a given time are not all alike and that their effect on the line profiles can be different when a *wave* is becoming visible on one limb and when it is leaving the visible hemisphere on the other one. The latter could happen if, for example, the velocity and/or temperature profile of a wave is asymmetric.

Figure 3-5 is an example of a more regular behavior observed in the C IV $\lambda\lambda$ 5802 and 5812 lines of ζ Pup (O4 If) during the course of one night. A preliminary analysis suggests that the bumps (most of the time there are two) which cross the line profiles are due to an $\ell = 4$, $m = -4$ mode with period 0.178 d. As can be seen, the (nearly) two cycles covered by the observations are not identical. This difference is caused by a slower additional variation which can probably be identified with an $\ell = 2$, $m = -2$ mode of period 0.356 d (see also Section VIII.B.2.b). If so, ζ Pup would be another example (see Baade, 1986, for references) of period commensurability with $P_m * m$ = constant for a given star. Another very puzzling observation of this star is that Balona (1986, private communication) found a photometric 1.2 d period which I have so far been unable to extract from my simultaneous spectroscopy and which clearly is not a simple alias of the spectroscopic 0.178 d period.

Other examples of O stars with absorption-line-profile variations most probably caused by NRP's are λ Cep (Penrod, 1986, private communication; O6ef), the O-type component of the Wolf-Rayet binary, γ^2 Vel (Baade and Schmutz, in preparation; O9 Ib), and HR 4908 (Baade, unpublished; O9 Ib). The detection of rapid radial-velocity variations in HD 108 by Vreux and Conti (1979) makes this Of star another candidate. A systematic spectroscopic survey of bright O stars for NRP's was recently begun by Fullerton et al. (1986, private communication). After the first 670 spectra of 33 program stars, they find NRP's in main-sequence stars to at least O7.5 in spectral type and a possible increase of the incidence of NRP's with luminosity. Their positive detections include HD 41161, HD 53975, and HD 93521.

(2) Properties and Implications of NRP's. From the number of the traveling features in the line profiles (times two to account for the invisible hemisphere) and the difference in arrival times between consecutive features (= period), one can determine the time which

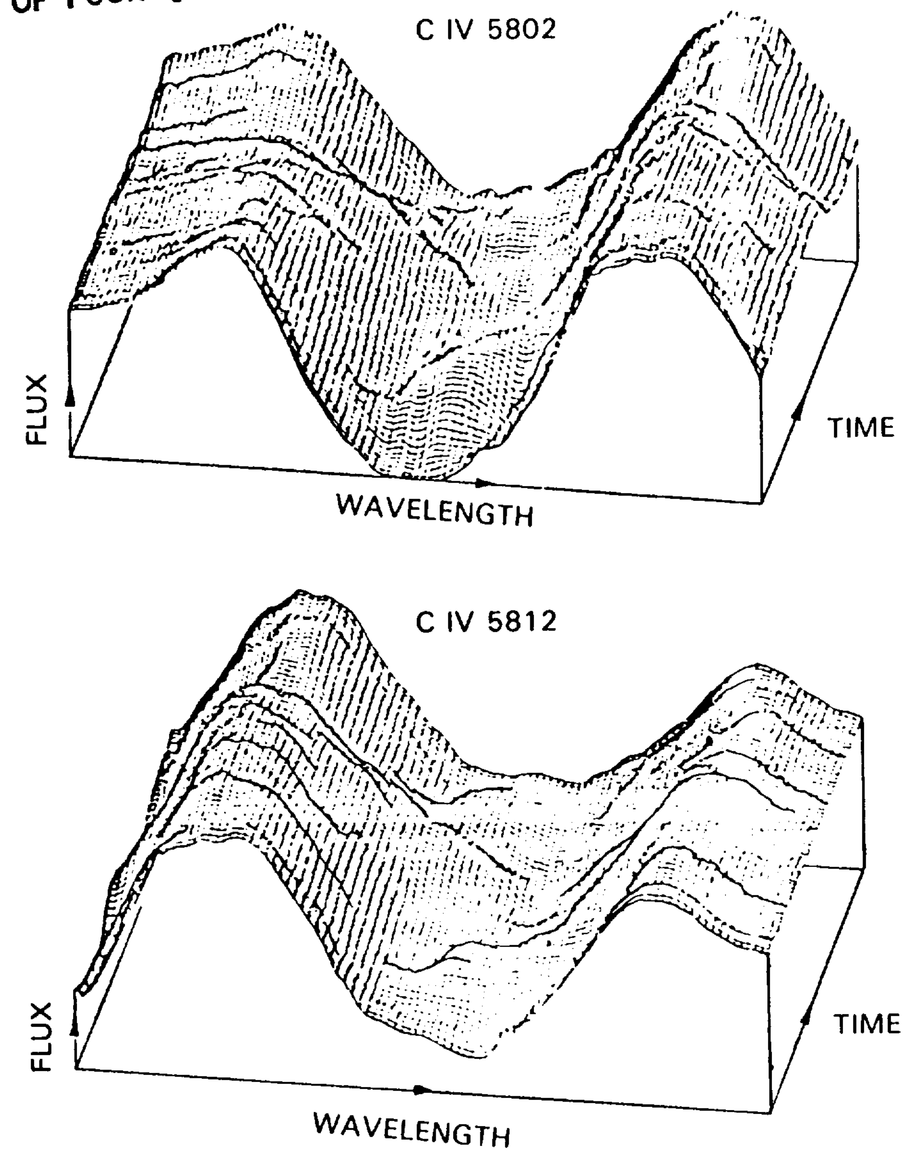


Figure 3-5. The variability of the C IV λ5802 (top) and λ5812 (bottom) lines in ζ Pup (O4 If). An uninterrupted series of 34 Reticon spectra obtained within nearly 8 hours at a mean time resolution of 14 minutes have been used to generate these plots. In addition to the two absorption lines, one emission component blueward of 5802 Å, one redward of 5812 Å, and a central one appearing in either plot are seen. These three discrete components are superposed on an extremely broad and shallow emission with a full width of at least 3000 km/s so that the true continuum is not seen in the original spectra. The range in flux shown is only 5 percent of the adjacent pseudocontinuum. Due to the limited resolution of the plot S/W used, the spectra appear to be smoothed noticeably. Note the similarity of the two plots, the narrow features which occur every 0.178 d in the blue wing and move across the line profile, the existence of additional slower variations, and the variability (approximately on an 8-hour time scale) of the emission components (most pronounced in the one shortward of λ5802) as an example of V/R variations. See Sections VIII.B.1.c(1) and (2) and VIII.B.2.b for discussion.

it takes the atmospheric pattern that gives rise to the spectral features to complete one revolution about the star. For ζ Pup, this time is 4 * 0.178 d = 0.712 d, whereas the rotation period is of the order of 5 days (see Moffat and Michaud, 1981). Hence, it appears to be inevitable to identify the implied high-propagation velocity with the phase velocity of traveling waves (i.e., NRP's). However, if one wants to exploit the pulsations to learn more about the stellar interior or to determine the effect the pulsations may have on the mass loss, a reliable mode classification is necessary. Unfortunately, as was demonstrated by Osaki (1985), not even detailed line-profile fitting leads always to a unique solution.

Simultaneous photometry is needed in addition to the spectroscopy in order to unambiguously detect temperature variations which may then be attributed to the compression and rarefaction of the atmosphere associated with a vertical velocity term. Because such simultaneous observations do not yet exist for pulsating O stars, little can be said about the pulsation modes at present. If, due to the restoring forces of, for example, pressure or gravity, the NRP velocity field has a radial component, rapid rotation would always add another restoring force. The eigenmodes are then neither p nor g modes nor Rossby waves, and their character will probably combine the properties of the two (Osaki, 1987b). Opposed to the p and g modes (also referred to as spheroidal modes) which are global eigenmodes of the star, Rossby (or quasi-toroidal) modes have no vertical velocity components at all and therefore do not transport information between different layers of the star. They are characterized by large eddies (cf. Saio, 1982).

There is as yet also no theoretical model that would predict a coupling of spheroidal modes with *m* commensurate periods. Such commensurability, however, is a natural property of quasi-toroidal modes whose frequencies are (to first order, Saio, 1982):

$$\sigma_0 = \frac{2m\Omega}{\ell\,(\,\ell+1)} \quad ,$$

where ℓ and m are mode degree and order, respectively, and Ω is the stellar rotation rate. This equation is not satisfied by the numbers given previously for ζ Pup, which, even though the first-order approximation in Ω may not be appropriate for this star, probably shows that the pulsation modes of ζ Pup are not pure Rossby waves. It has also been pointed out (Penrod, 1986, private communication) that, in rapidly rotating stars, approximately m commensurate periods will be observed if, in the corotating frame, the two velocity patterns propagate very slowly so that the mode-coupling is only an illusion caused by the much faster rotation. In ζ Pup, however, the rotational angular velocity is much *smaller* than the pulsation phase velocity in the inertial frame, and if period commensurability is found in the observer's frame, it must also exist in the corotating frame. Longer time series of observations are currently being analyzed to check this possibility.

The supergiants, ζ Pup, HR 4908, the one in the γ^2 Vel system, and several others of spectral type B (Baade and Ferlet, 1984; Baade, 1984b, and unpublished data), all have rather broad lines for their luminosity class. In most of them, intermediate- to high-order modes, which are less easily detected in narrow-lined stars, appear to dominate, and one may wonder whether this explains the failure until now to detect NRP's in OB supergiants with low *V sin i*. The low spatial frequency line-profile variations of ζ Pup are in conflict with the complete nondetection of NRP's in narrow-lined O supergiants, but a resolution to this problem can only come from a larger survey. If the waves observed at the surface in all cases directly reflect the pulsation of the entire star, a negative result could throw much light on the role of (differential) rotation for the excitation (Ando, 1985) and perhaps for mode selection (Baade and Ferlet, 1984) of NRP's. Section VIII.C will describe another possible difference between broad- and narrow-lined supergiants which may mean that the two groups of stars in fact do differ in their pulsation properties. (Tentatively, one could place the limit in *V sin i* between the two groups somewhere between 100 and 150 km/s.) As outlined previously, theory also predicts a strong effect of rapid rotation on the eigenmodes. Finally, it deserves mentioning that I have not yet found a broad-lined OB supergiant that does not show the signature of intermediate- to high-order nonradial pulsation modes.

The higher order modes detected thus far all have periods of a few hours and are therefore not to be confused with the g-modes which have been suggested as an explanation of the usually much slower photometric variability of early-type supergiants. In principle, however, they may nevertheless be responsible for those much slower variations if the latter reflect the beating of many nonradial modes which each have shorter periods. Depending on the number of modes and the stability of their amplitudes, the beat periods would only appear as quasi-periods, and the typical profile variations of single modes would no longer be recognizable (cf. the work by Smith and Ebbets, 1981, on ρ Leo, B1 Ib). This alternative explanation evidently raises the question why, then, such a situation does not prevail in broad-lined supergiants (cf. the work by Baade and Ferlet, 1984, on γ Ara, B1 Ib).

Lucy (1976) has pointed out that multimode nonradial pulsation can observationally become indistinguishable from macroturbulence. The dependence of macroturbulence on spectral type and luminosity class found by Slettebak (1956) and Conti and Ebbets (1977) may therefore become interesting for the investigation of NRP's in OB stars. But equating the lack of narrow-lined O supergiants with line-broadening by multimode NRP's is not straightforward because the line-broadening effects of NRP's also depend on the inclination angle. This dependence is very different for different types of modes, and electron scattering must be considered as a possible additional line-broadening mechanism. In any case, explanations using a multimode pulsation are clearly unsatisfactory when they rely on something unobservable, and they must not discourage observers from searching for alternatives.

2. Variations in the Low-Velocity Regime of the Wind. In the current small sample of spectroscopically confirmed nonradially pulsating O stars, stars with known emission lines are either supergiants or Oe stars. This section therefore deals only with these two subcategories.

Section VIII.B.1.a has questioned whether NRP's are more likely to exist in stars with rotationally broadened absorption lines (higher order) or whether they are only more easily detectable. The emission lines, if any, of such stars are also broad, and their appearance generally is the one of emission from a rapidly rotating disk. Conversely, the P Cygni character of emission-line profiles is usually more prominent in stars with narrower lines. The following paragraphs describe a tendency of these two basic types of emission-line profiles *in supergiants* to also exhibit different patterns of variability. If a correlation with the pulsation characteristics is also borne out by future observations, differences regarding emission lines and their variations between fast and slowly rotating supergiants would provide an additional, indirect indication of a general difference in their pulsation properties. The reader is warned that this distinction is based on observations of only a few stars (but augmented and supported by observations of B stars not discussed here) and that transition types and even counter examples are likely to exist. The reader is explicitly reminded of the uncertainty regarding the apparent nondetection of nonradial pulsations in narrow-lined supergiants (Section VIII.B.1.c).

a. Stars Without Clear Evidence of Nonradial Pulsations. Since the work of Rosendahl (1973) and Ebbets (1979a, 1982), it is well known that the Hα emission of OB supergiants is variable. Ebbets's observations also give some examples of considerable variations within a few days. Extensive monitoring of a small number of bright OB supergiants (Baade and Ferlet, in preparation) shows more specifically that the emission of most stars varies cyclically with typical time scales of 5 to 15 days. Changes within one night can be detected but were, in all stars observed, only part of that slower variability (but see also Section VIII.B.2.b). Bright supergiants of spectral type B1 or earlier which were found (Baade and Ferlet, in preparation) to show the behavior described below include κ Ori, ϵ Ori, μ Nor, HR 6334, and perhaps τ CMa and ζ Ori (the classification of which is based on the profiles in Ebbets (1982)), but the Hα variability of the B8 Ia supergiants, β Ori and μ Sgr, is qualitatively the same.

A cycle (e.g., Figure 3-6a) usually appears to begin with a relatively abrupt increase of the emission's equivalent width by as much as a factor of 2. The subsequent decline of the emission is generally slower and accompanied by pronounced profile changes, implying changes in the velocity profile of the wind. Whether the equivalent-width changes mean that the mass-loss rates are also variable can only be determined from a detailed modeling. As can be seen from the work of Olson and Ebbets (1981), simple spherically symmetric models using the Sobolev approximation cannot fit well the line centers, where most of the variability is. However, the fact that, in stars with strong emission wings, these wings are rather constant (cf. Ebbets, 1980a; Olson and Ebbets, 1981), should mean that at least the mass-loss rate averaged over approximately 1 day (the time that it takes standard velocity laws to accelerate matter to near-terminal velocities) is fairly constant (i.e., no *major* mass-loss-rate variations would be associated with this cyclic variability in Hα (see also Section VIII.B.3)). A reasonable qualitative interpretation is perhaps that, in each cycle, some matter is lifted above the star and then slowly dispersed and accelerated in the wind; some fraction may also return to the star (cf. end of this section).

Observations in different seasons and comparison with profiles in the literature further show that the variations of each star follow an individual characteristic pattern so that two stars may not only differ in instantaneous Hα profiles, but their respective cycles may not have a single phase in common. The time scales are similar to the photometric ones, and it will be very interesting to make simultaneous observations to determine if the two phenomena are in fact only two aspects of one process. There is now only little to no evidence of a correlation between changes of the Hα emission and variations of photospheric line profiles (cf. Figure 3-6), even though the time scales are again comparable.

Hot stars near the Eddington limit (S Dor-type or luminous blue variables) sometimes show outburst-like drastic changes of their entire spectrum which probably indicate strongly increased mass-loss rates (Wolf, 1986; Wolf and Stahl, 1983; Walborn, 1984c). These stars are important because of their possible evolutionary connection with even more extreme objects like η Car (cf. Lamers, 1987) and with Ofpe/WN9 stars (Stahl, 1986). With future observations, it might be interesting to determine if the Hα variability of less luminous OB supergiants described previously can be called an enormously scaled-down version of S Dor-type outbursts.

An excellent tracer of circumstellar matter within few stellar radii is the polarization controlled by electron scattering. Most OB supergiants show some intrinsic polarization which, moreover, is variable on a time scale of days in those stars that were observed repeatedly (Hayes, 1975, 1978, 1984; Snow and Hayes, 1978). The changes seen represent ordered large-scale variations in the distribution of the near circumstellar medium because the daily measurements can usually be joined by a steady line in the Stokes Q,U plane. On the other hand, the path so described is reminiscent of Brownian motion.

The difference between the rather irregular variations of the mass-loss geometry implied by the polarimetry and the more pronounced cyclic behavior of Hα may perhaps be seen as evidence that the Hα activity is not strongly localized (e.g., due to magnetic fields). This also supports the impression of a more event-like process. Global mass-loss events are difficult to accomplish with a single low-amplitude, low-order NRP mode alone. The beating of many modes with shorter periods may again

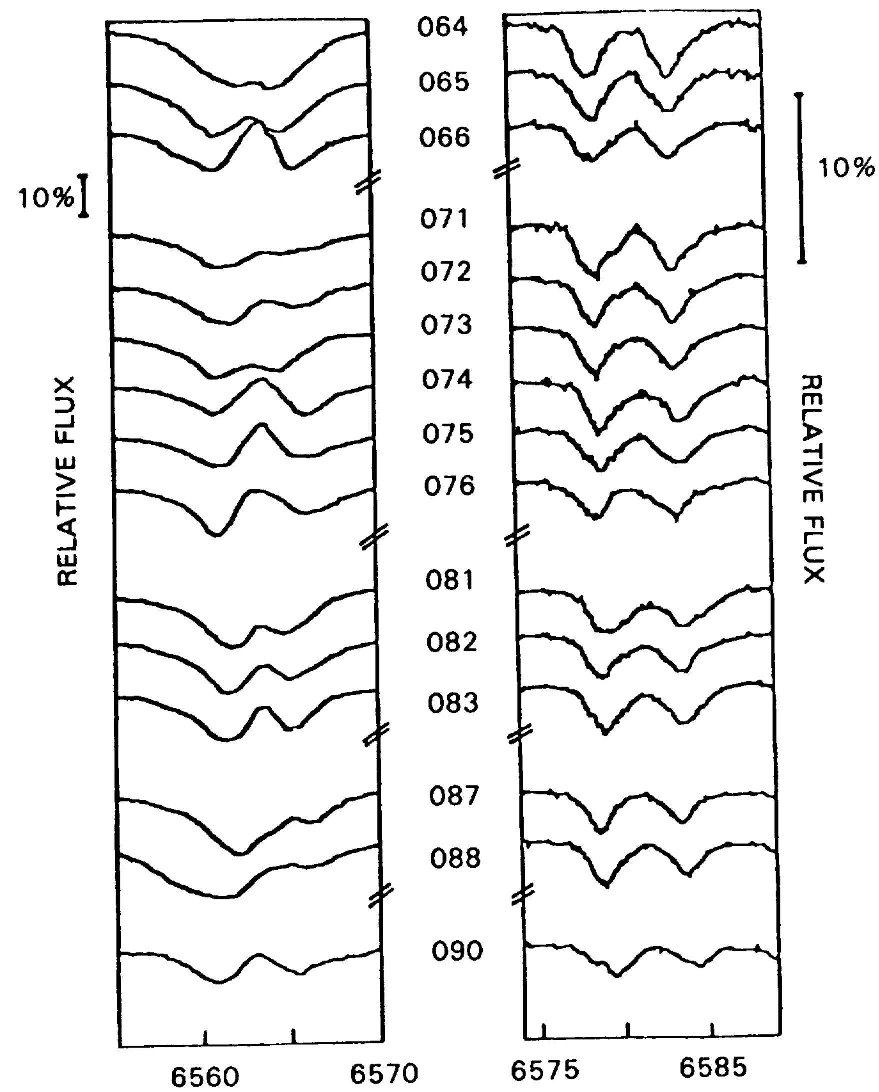


Figure 3-6. (a) An example of cyclic variability of the Hα P Cygni profile in an OB supergiant (κ Ori, B0.5 Ia) without very broad lines. Although the mean emission equivalent width is smaller than that observed by Ebbets (1982), the variations are similar and therefore mark a stable, but clearly not singly periodic, long-term phenomenon. Observing dates are labeled. Note the repeated sudden appearance of a comparatively strong and narrow central emission peak, the strong depression of the blue wing of some profiles, and the pronounced variability of the red absorption component (which in Ebbets's data was sometimes stronger than the blue one). The far wings of Hα are not included in this presentation. (b) Simultaneous profiles of the C II λλ 6678 and 6683 lines. Note the presumably cyclic behavior of these lines also, the permanently depressed blue wings, and the lack of any correlation with the variations in Hα. The profile variations are not dissimilar to the effects of low-order nonradial pulsations, but a confirmation is needed from observations of other lines at higher time resolution. In both panels, strong telluric lines have been removed from the profiles (especially λλ 6683), and the flux scales of the two panels are different (see bars). In 2 nights with multiple observations, only very small or no variations were detected in all three lines.

be a way out (cf. Section VIII.B.1.b). Also noteworthy in this connection is the marked contrast to the polarization variations seen in pulsating Be stars like ω Ori (Hayes and Guinan, 1984) where the polarization angle is constant irrespective of the degree of polarization, which shows some correlation with the Hα emission. The implied increased mass-loss rate in a fixed plane is consistent with both the fast rotation of Be stars and the equatorial amplitude maxima of many types of NRP modes.

Lastly, some of the stars in this group occasionally show absorption components longward of the emission peak in Hα which are deeper than the *normal* P Cygni absorption. This was first noted in $\varkappa$ Ori by Olson and Ebbets (1981), who conclude that it cannot occur in rotating or spherically symmetric expanding winds and may require infalling matter. One way to think of this is as the motion of gas blobs along (closed) magnetic field lines. Underhill and Fahey (1984) have even suggested that much of the mass loss is caused by magnetic activity. Because β Ori is one of the most extreme examples of this kind of variability and one of the few nonchemically peculiar early-type stars with reportedly significant magnetic field measurements (Severny, 1970), this hypothesis should be accessible to direct observational tests. Variations found in quasi-simultaneous B-band polarimetry (Hayes, 1986) and Hα spectroscopy (Baade and Ferlet, in preparation) were uncorrelated.

b. Known Nonradially Pulsating Stars. Closer inspection of Figure 3-5 shows that the emission components straddling the two C IV absorption lines in ζ Pup vary in strength. The time scale of that variation is not the short period of the $\ell = 4$ mode, but is perhaps related to the suspected (Section VIII.B.1.c(1)) slower pulsation mode. A series of 34 Hα profiles obtained over 10 nights in January 1986 (Baade and Gathier, in preparation) exhibits a similar V/R-type variability. Again, the time scale is compatible with that of the pulsation, but with an average of only three observations per night and superposed nightly equivalent-width variations, a periodicity is not readily deduced. Nevertheless, the much shorter time scale and the occurrence of the phenomenon in many lines are additional criteria to distinguish the V/R variations from the Hα variability described in Section VIII.B.2.a. Furthermore, if the latter type of variability occurs in addition to V/R variations, its amplitude must be significantly smaller than that in narrow-lined stars. The nightly equivalent-width variations seen in ζ Pup are presumably one example. Another O supergiant which appears to combine the two types of variations in a similar way is α Cam (O9.5 Ia; Ebbets, 1980a); incidentally, Smith (1986b) lists it as a nonradially pulsating star.

Much more regular rapid V/R variations have been seen in many Be stars (Smith and Penrod, 1985; Penrod, 1986; Baade, 1984a, and unpublished data), where the presence of emission components, sometimes in numerous non-hydrogen lines is probably linked to phases of enhanced mass loss. A low-order ($\ell \approx 2$) mode appears to be another prerequisite, and the V/R variations have the same period as the pulsation.

Further examples of V/R variability in O stars are λ Cep (see below) and the O9 Ib star, γ^2 Vel. λ Cep appears to be especially interesting because Hayes (1978) has found significant night-to-night polarization variations in the B-band. Unfortunately, since the interstellar polarization is undetermined but presumably large compared to the amplitude of the stellar variability, it is not known whether, as in some Be stars (cf. Hayes and Guinan, 1984), the near constancy of the measured polarization angle also implies constancy of the intrinsic polarization angle. This would be expected if NRP and/or rotation impose some fixed geometry on the stellar wind structure. Observations by Lupie and Nordsieck (1983) show that this is at least not a general property of O stars.

I have not seen major changes in the V/R variations of ζ Pup within 1 month or between 2 consecutive years; in fact, they seem to persist over years since the Hα variability reported by Wegner and Snow (1978) is probably the same. A reanalysis of Moffat and Michaud's (1981) Hα data of ζ Pup beyond its Nyquist frequency yields the strongest peak in the power spectrum at 0.356 d (i.e., twice the period of the $\ell = 4$ mode in 1985 and 1986). 0.356 d may also be the period of the presumed low-order mode. The intensity change found over 2 years by Conti and Niemela (1976) and interpreted by these authors as a change in the mass-loss rate by $\sim$20 percent is not much larger than the much more rapid variations seen in January 1986. Observations of the He II λ4686 line (Conti and Frost, 1974; Hutchings and Sanyal, 1976; Leep and Conti, 1979; Grady et al., 1983) also document the long-term existence of rapid V/R variations in λ Cep. Because the emission is visible only in the line wings, the line-emitting region is a rotationally flattened structure (cf. Section VIII.B.2.a for related polarimetric observations of Be stars and their interpretation) and does not form a uniform disk or ring, but because of the variability with the pulsation period, must somehow be subdivided into m azimuthal sectors, where m is the azimuthal quantum number of the pulsation mode. (This has thus far been observed for only low-order $m \approx 2$ modes.)

NRP's evidently have a very pronounced effect on the wind at its base, but the detailed process which leads to the variable line emission has not yet been identified. There are at least three possibilities: (1) a change of the ionization balance caused by temperature variations associated with the radial component of the pulsation, (2) variable amounts of extra gas tossed up as a consequence of the pulsational velocity field, and (3) amplification of the pulsation amplitude to shocks in the tenuous outer atmosphere. The variability of the V/R amplitude is much larger than any variations in the nonradial pulsation amplitude. Therefore, either this difference is due to a nonlinear coupling with other, otherwise undetected modes or the pulsations provide a direct means of probing the wind at its base.

A related and perhaps even more significant observation is that the absorption component in the Hα profile of ζ Pup shows essentially the same type of variations as that of the photospheric lines. The trouble is that this Hα feature is blue-shifted by about 100 km/s and is therefore clearly formed in the wind. The implied strength and coherence of the coupling between wind and photosphere is rather surprising, but I have not yet found a different interpretation. The visibility of these relatively low spatial frequency features may also throw some additional light on the extreme broadening of the emission which often is attributed to electron scattering. (Ebbets (1980b) and Mihalas and Conti (1980) also consider corotation of the wind out to a few stellar radii due to a weak global magnetic field, but Barker et al. (1981) could not detect a magnetic field in ζ Pup.)

Like Be stars, Oe stars also suffer major variations of their circumstellar envelopes (Vogt and Penrod, 1983; Divan et al., 1983) which are caused by a (sudden) increase of the mass-loss rate. There are indications of a link between such major mass ejections and changes of the pulsation amplitudes. Attention to the probable existence of such a correlation was drawn first by Bolton (1982) for the B2 IVe star, λ Eri, and by Vogt and Penrod (1983) for ζ Oph. It was later directly observed in λ Eri by Penrod (Smith and Penrod, 1985) and can perhaps also be traced back in observations of the B6 IIIe star, o And, over decades (Baade, 1985, and references therein). Ando (1986), Castor (1987), Osaki (1986b), and Willson (1986) have carried out theoretical studies of how nonradial pulsation can lead to mass loss. At least some of the proposed mechanisms appear to be applicable to only a limited range in the frequency of mass-loss events, whereas recent observations indicate that there is probably a nearly continuous spectrum from a few major to many small events, and of course one would like NRP's to explain the entire range (if any). However, the direct observational proof of such a causality may be difficult to obtain because the observed decrease of the pulsation amplitude during phases of enhanced mass loss (Vogt and Penrod, 1983; Smith and Penrod, 1985) alone is not sufficient as such. If there is any other (as yet unknown) process that can eject considerable amounts of mass, it seems equally conceivable that it can also temporarily damp the pulsation or destroy the coherence of its velocity field.

3. Possible Connections to the Variability of the Outer Wind. Variability is not a prominent property of OB-star winds as far as temporal changes of the mass-loss *rates* are concerned (Prinja and Howarth, 1986) although Oe/Be stars will probably have to be excepted. However, variations do not have to be time-dependent. Observations of the mass loss from hot stars generally agree well with the line-driven wind theory (Abbott, 1984), but the scatter of mass-loss rates about the mean value of a given location in the HR diagram is larger than can be explained by observational errors, binarity, etc. Garmany and Conti (1984) found no dependence of the residuals from the mean $L - M_{bol}$ relation on static parameters like mass, radius, or chemical composition. A search for a correlation with any stellar variability is not yet possible because too few stars have been studied. However, the large apparent range in variability amplitudes, in particular, lets this hypothesis appear to be worth pursuing further.

Variability of the outer wind is therefore nearly synonymous with the variability of the high-velocity discrete absorption lines which are the subject of a different chapter of this volume. Although they are usually attributed to variable mass transfer in interacting binaries (often with a compact component), however, X-ray and nonthermal radio-flux variations may also be intrinsic phenomena. The question arises whether any of these variations seen in the outer wind can be traced back to some variability deeper in the stellar atmosphere.

An intriguing observation in conjunction with the narrow UV components is that they also occur in stars whose luminosity is so low that stellar-wind theory does not make a clear prediction as to whether or not they have a self-initiated line-driven wind (Abbott, 1984). Hence, narrow components are not necessarily a property of line-driven winds as such. Six additional observations—(1) Henrichs' (1984, see also this volume) identification of *all* these stars with Be stars, (2) reports by Penrod (Smith and Penrod, 1985) and Baade (1986) that Be stars differ from Bn stars in that the latter do not usually display a low-order NRP mode, (3) the connection of these low-order modes to the V/R-type variability of emission components (Section VIII.B.2.b), (4) the reduced amplitude of low-order modes during major Be-star outbursts (Penrod, 1986), (5) the equatorial amplitude maximum of the suggested sectorial modes, and (6) the nondetection of narrow components in Be stars with $V \sin i \leqslant 150$ km/s (Henrichs, 1984)—appear to combine as persuasive evidence in support of the stellar variability hypothesis for the narrow components in general and models involving density enhancements such as shells (= roughly spheroidal expanding envelopes of the entire star) or *puffs* (= smaller mass parcels with an *a priori* arbitrary spatial distribution) in particular. (See Henrichs' contribution to this volume for references and further discussion.) Also important is the simultaneous high incidence of NRP's and variable photospheric line profiles (Fullerton, Gies, and Bolton, 1986, private communication) and the ubiquity of discrete UV components in O stars. With regard to the latter, in a sample of nearly 200 O stars, Prinja and Howarth (1986, private communication) find them in virtually every star with well-developed but unsaturated P Cygni profiles. Finally, both phenomena appear to be more pronounced in higher luminosity classes.

In spite of several efforts, however, observers have not succeeded in unambiguously documenting for any particular star the propagation of a perturbation in the wind from the photosphere to the narrow-component-forming regions. In λ Cep, Grady et al. (1983) saw variations of the sharp blue edges of the absorption portion of several UV P Cygni profiles simultaneously with the usual V/R-type variability in the emission of He II λ4686. Similarly, Wegner and Snow (1978) detected in the Hα emission of ζ Pup what appears as its V/R variations (Section VIII.B.2.b) and a possibly similar variability in the He II λ4686 emission, whereas the overall appearance of most wind features in the UV was constant. However, Wegner and Snow also report "simple excursions from the *quiescent* profile intensity over several angstroms." The time scale appears to agree well with the known pulsation periods. Recent IUE spectra (Prinja and Howarth 1986, private communication) obtained simultaneously with some of the ground-based observations of this star described previously (Sections VIII.B.1.c.(1) and VIII.B.2.b) detected differences only between wind profiles from different days. Compared with the optical spectral range, however, the attainable S/N is a much stronger limiting factor in the UV (with IUE more than with Copernicus). With a better photometric accuracy, rather different conclusions might be reached. On the other hand, Snow and Hayes (1978) described UV line-profile changes of δ Ori A (a double-lined binary; Baade and Ferlet, in preparation) as small outbursts, but could not detect variable polarization at the 20-percent significance level. (In January 1985, this star did not show a trace of emission in Hα (Baade and Ferlet, in preparation).)

There are basically two different ways to resolve this apparent conflict. One requires model calculations to determine if, for example, NRP-induced perturbations can maintain their identity throughout the wind. Depending on the dominance of their intrinsic instabilities (Lucy, 1984b; Owocki and Rybicki, 1985), stellar winds are not necessarily a good causality-preserving medium. Radiation-driven shocks, alone or in conjunction with pulsation-driven shocks, may also impose large-scale structures on the wind.

On the observational side, a loophole may have been the limited time resolution and/or coverage of the simultaneous multifrequency observations described previously. As discussed by Henrichs (1984 and this volume), one of the main obstacles against the density-enhancement hypothesis for the narrow components, namely that they are hardly ever seen at low velocities (but note the spectacular counter example observed by Prinja and Howarth (1986) as shown by Henrichs in this volume), weakens with increasing lifetime of the features if the acceleration time remains short, so that the probability of their detection at low velocities would decrease in proportion. Model calculations by Prinja and Howarth (1985) show that, depending mainly on the acceleration and radial-density profile of an assumed shell, the lifetime of a narrow component produced by a single shell could reach a few days. With insufficient temporal sampling, observations of stars which eject shells (or big puffs) at intervals similar to the lifetime of the narrow components could lead to the impression of narrow components that are variable in strength and shape, but stay at roughly the same velocity.

If the V/R variability associated with NRP's actually leads also to the formation of little puffs, each of them would not become prominent at low velocities because of the rapid initial acceleration (cf. Henrichs' Figure 4-48 in this volume). At the excellent S/N attained with the Copernicus satellite, the phenomenon may, however, have been detected by Gry et al. (1984) as transient features in several Lyman lines. Farther out in the wind, all puffs would reach roughly identical terminal velocities and therefore add up to a seemingly single and much stronger feature. However, this latter effect does not occur in models calculated by Scuflaire and Vreux (1987) for Wolf-Rayet stars.

The same picture clearly becomes questionable in the case of stars without evidence of NRP's. Stretching Prinja and Howarth's (1985) model to its limits, the time scales of the Hα variability in these stars (Section VIII.B.2.a) are perhaps just short enough to also permit the formulation of a similar hypothesis for these stars if the variable Hα emission also means variable mass-loss rates. Cassinelli et al. (1983) have obtained simultaneous X-ray and UV observations of ϵ Ori (B0 Ia) and χ Ori (B0.5 Ia), which belong to this category. In neither star were significant variations seen over a 12-hour interval, which is consistent with the one order of magnitude longer time scale found in ground-based observations (cf. Figure 3-6). The mean Hα emission in this type of OB supergiants undergoes additional long-term variations. If they are caused by density variations, they may be related to the increase by nearly 50 percent over 1 year in the X-ray flux of κ Ori (Cassinelli et al., 1983) since it is, for example, conceivable that, at the interface between a high-density structure which is more slowly accelerated than the average velocity law and the ambient much faster wind, a shock front develops.

With the foregoing inclusion of more detailed ground-based observations into the density enhancement hypothesis for the narrow components, observational tests can be more stringently focused. Apart from stars showing all the variations mentioned (they are probably many), stars like δ Ori A (which is a binary) deserve attention as they have displayed variable narrow UV components but, at least at other times, no Hα emission. One would also expect that single-mode NRP leads to the formation of puffs rather than shells (defined to extend over a complete sphere) so that the resulting narrow components are not totally black. Intriguing test objects are stars like HD 153919 (= 4U 1700-37; O7f), which are orbited by a compact object whose variable accretion rate and X-ray flux (see Pietsch et al., 1980) may be used to probe the (variable) structure of the primary's wind. After having eliminated the proposed black hole from the X-ray binary, X Per, Penrod and Vogt (1985) pointed out that the 22.4-hour modulation of the X-ray flux and its apparent lack of strict periodicity are plausibly explained by the assumption that the O9e star is a nonradial pulsator with that period and that its wind, and therefore the accretion onto the neutron star companion, are modulated by the pulsation.

If the perturbation of the wind by the star also leads to shocks (conditions of their formation are discussed by Castor (1987)) comparable to the intrinsic instabilities considered by Lucy and White (1980), there should also be variations on short time scales of the X-ray flux from single O stars. Because of the very low count rates not exceeding 1 ct/s with modern equipment, significant results will be difficult to obtain. However, a change by a factor of 2 within 5 days seen by Snow et al. (1981) in 15 Mon (O7 V((f))) demonstrates that some detectable rapid variations exist. Similarly, within White's (1985) model for the variable nonthermal radio emission from O-type stars (Abbott et al., 1985, and references therein), changes in the stellar variability on a time scale of 2 weeks may also become detectable at radio frequencies.

C. Wolf-Rayet Stars

At least at first glance, the optical thickness of the winds of most Wolf-Rayet stars does not make the search for intrinsic variability of the underlying stars very promising. It would be all the more important to recognize variations as stellar in origin since this would permit, however hazily, a direct glimpse of the star. For a long time, moreover, almost any apparently periodic variation was routinely fed into binary models. On the other hand, photospheric velocity fields, notably NRP's, have played some role in the intensive theoretical discussion about the origin of the high mass-loss rates of Wolf-Rayet stars and the momentum of their winds, which exceeds the single-scattering limit (cf. Chapter 4, Section IV).

1. Long-Term Flux Variations. Infrared brightenings by typically 2 magnitudes in the K-band, which can be interpreted as the formation of an optically thick dust shell with a temperature of about 1000 K, have been observed in HD 193793 (= WR140; WC7) in 1970, 1977, and 1985 (Hackwell et al., 1976, 1979; Williams et al., 1978, 1986). The star is probably not a short-period binary (Conti and Roussel-Dupree, 1981, as quoted by Fitzpatrick et al., 1982), but Williams et al. (1986) report that the apparent 7.9-year period can also be recovered from published absorption-line radial velocities. In a fairly eccentric orbit ($e=0.7$ to 0.8), then, the dust formation could be the result of the variable interaction between the winds of the Wolf-Rayet star and its O-type companion. The IR variations have not been detected by contemporaneous *UBV* photometry (Fernie, 1978), and at least 2 years after maximum had not caused the appearance of molecular absorption lines detectable with the IUE satellite (Fitzpatrick et al., 1982). The inferred dust shell therefore does not intercept the line of sight (i.e., it is not spherically symmetric, which could further support the binary hypothesis).

On the other hand, a similar IR brightening has been observed in a newly discovered WC9 star, but in a low-resolution, high S/N spectrum, no absorption lines suggestive of an O-type companion were found between 5800 and 11000 Å (Danks et al., 1983). Hackwell et al. (1976) also reported a temporary flux increase in the K-band by 0^m7 for HD 192641 (= WR137; WC7). Like HD 193793, this star's spectrum shows absorption lines but no evidence of orbital motion with $K > 20$ km/s and $P > 2$ years (Massey et al., 1981). Finally, the fadings of HD 164270 (= WR103; WC9) by $\sim 1^m$ in visual magnitude in 1909 and 1980 (Massey et al., 1984) may have been similar events if the dust cloud partly obscured the star, although the lack of color variations would impose rather stringent constraints on the properties of the dust. This star also shows short-term radial-velocity and photometric variations (see Vreux, 1984, for references).

Although a binary model is the most promising for HD 193793, the frequency of such eccentric long-period binaries among WCL stars would have to be rather high if one explanation is to account for all four examples given. In view of the insufficient temporal coverage of the observations, the intrinsic-variability hypothesis still appears defensible (see also Section VIII.C.4). In either case, it would be important to distinguish between a variable condensation of dust in an otherwise constant wind and a variable mass-loss rate. With estimates of 1.0×10^{-6} to $1.5 \times 10^{-4}\ M_\odot$ (Hackwell et al., 1979), the total mass involved could be comparable to the mean annual mass loss, which, by mere grain formation at unchanged mass-loss rate, may be difficult to accomplish.

2. Variations on Short Time Scales. The intention to discriminate between single stars and binaries and the knowledge that orbital periods would not normally be shorter than 1 day have clearly governed the typical observing frequencies and probably were not without influence on the interpretation of the results. To check the latter effect and as precursor to observations not limited by the former, Vreux (1985) has undertaken a survey of observations and periodograms published of stars sometimes suspected to have a compact companion in order to search for indications of alternative short ($\leqslant 1$ d) periods. For the assessment of the impact of binarity on the evolution of Wolf-Rayet stars, the unambiguous detection of compact companions is very important, but is made difficult by the general lack of significant excess X-ray radiation so that one must rely on more indirect indicators. However, the combination of single-linedness, variability on a relatively short time scale, and small amplitude also makes these stars prime candidates of being intrinsically variable.

For four of the 12 stars in his sample, Vreux (1985) finds that observers have mentioned alias periods below 1 day. All these periods were rejected by the original authors for various reasons, but in most cases, the exclusion of such periods on purely mathematical grounds would probably require observations at higher time resolution. Until such observations are available, Vreux's work is the strongest hint thus far at the possibility that many Wolf-Rayet stars may be periodic intrinsic variables. This variability should not be limited to single stars, and in his analysis of the residuals from the orbital radial-velocity curve of the well-known binary, HD 90657 (= WR21; WN4 + O4–6), Vreux finds regular variability with a period of 0.44 d. On the other hand, recent surveys show that the optical brightness (Moffat et al., 1987a) and polarization (Moffat et al., 1987b) of some

Wolf-Rayet stars is almost perfectly constant on short-to-medium time scales, whereas it is mostly binaries which display periodic variations and significant amplitudes. But there are also additional, apparently random variations in binaries. The amplitude of this intrinsic noise appears to peak at spectral type WN8.

During his search for reports on short-term variability, Vreux (1985) also noted a curious algebraic scheme of a few equations that approximately relates the frequencies, harmonics, and/or alias periods of all stars of his sample except one. He conjectures that, in the case of intrinsic variations such as nonradial pulsation, this apparent homogeneity of Wolf-Rayet stars could be related to the similar small scatter of their mass-loss rates, possibly because the pulsations contribute significantly to the mass loss. One may add that, if the variations are also caused by traveling waves (in the case of very long NRP periods in the corotating frame only; cf. Section VIII.B.1.c (1)), the rotation rates of Wolf-Rayet stars would have to be similar because the observed periods still include the effect of the stellar rotation. Vreux continues to compare the occurrence of low-order harmonics in the proposed frequency scheme with reports on NRP mode-switching involving period changes by about a factor of 2 in the 53 Per stars (cf. Smith, 1981). In other OB stars, however, no clear-cut case in which one mode gained in amplitude at the expense of another has been found (see Baade, 1986), and a reinvestigation of the phenomenon in the 53 Per stars could be useful. Nevertheless, the possible implication of at least a very homogeneous subgroup of Wolf-Rayet stars is intriguing and should be further pursued. At present, formal objections are that the sample of 11 stars has been split into six groups to accommodate all observed periods and that there is no evident reason to prefer 1–f to f + 1 aliases as in Vreux's (1985) Table I, especially not when studying the possibility of rapid intrinsic variations.

In only two cases have observers considered intrinsic variability as a serious alternative to the default binary model. These can serve as an illustration of the uncertainties likely to be encountered in other stars. Vreux et al. (1985) find convincing evidence that, in HD 192163 (= WR136; WN6), the radial velocity of the emission lines due to N IV λ4058 and N III λ4100 vary periodically with P = 0.45 d or 0.31 d. In a binary model, the orbit of the implied compact companion is deep in the Wolf-Rayet atmosphere, and the system is seen in an advanced stage of spiraling-in of the companion. The drag by the Wolf-Rayet atmosphere should have circularized the orbit, but Vreux et al. find an eccentricity of 0.25 and therefore discard a binary model. On the other hand, a stellar atmosphere with an embedded compact object must be substantially perturbed, and it requires a detailed knowledge of this complicated situation to exclude the possibility that the eccentricity is spurious as in many less exotic interacting binary systems.

For HD 96548 (= WR40; WN8), Smith et al. (1985) find a relatively complicated power spectrum of its photometric continuum variations, but are able to show that it was stable over several years. They deduce 5.879 d and its 1 c/d aliases, 1.204 and 0.855 d, as the most probable periods and report that, in data strings with sufficiently high time resolution, the 1.204 d period may be preferable. A fairly strong peak at 1.831 c/d, a 1 c/d alias of the 1.2 d period, may also support the 1.2 d period, which is about 1/4 of an earlier period determination (see Smith et al., 1985). Although the discussion of a 5.9 d binary with compact component does not yield a very convincing solution, Smith et al. nevertheless find it difficult to opt for the shorter period because the IUE satellite observations showed no variations within 2 days, but they also note that the IUE data are similarly incompatible with a 5.9 d modulation.

3. Ultrarapid Variations. Occasionally ultrarapid variations (i.e., within a few minutes) are reported. They are not reviewed here because there is no plausible way of interpreting them as intrinsic to a Population I Wolf-Rayet star and their reality has not been established beyond doubt. As to the second point, one might, for example, add to the bibliography of such studies compiled by Jeffers et al. (1985) the negative results of the coordinated photometry, polarimetry, and spectroscopy by Haefner et al. (1977), who found no agreement between power spectra of simultaneous data strings and note that the enormous brightness of a star like γ^2 Vel results in the measuring accuracy not being photon-noise-limited, but governed by various atmospheric and instrumental effects (cf. Lacy, 1977).

4. Related Observations and Discussion of the NRP Hypothesis. Smith et al. (1985) compare the behavior of HD 96548 with the apparently random variations in the UV spectra of (single) O stars, but reject this comparison as inappropriate because periodic variability in the visible, like the one of the Wolf-Rayet star, was not known at that time as a fairly frequent property of (single) O stars. With the evidence accumulated since of NRP's in O stars, the comparison with O stars would now probably end differently. This leads directly to the question of what observable effects can be expected if Wolf-Rayet stars are in fact nonradial pulsators as considered by Smith et al. (1985), Vreux (1985), and Vreux et al. (1985). Because in Wolf-Rayet stars with optically thick winds, most of the information contained in the radiation emerging from the central star is reprocessed, it is clear that the classical criteria for NRP's (cf. Lesh and Aizenman, 1978), like color-to-light and radial velocity-to-light amplitude ratios and phase differences, are probably not applicable. Even the time scales may not be those of the central stars but of the envelopes in which the reprocessing takes place and which could at least play the role of a filter.

Maeder (1985) gives a range of 15 to 60 minutes for the period of the radial fundamental pulsation mode of Wolf-Rayet stars so that the observed time scales clearly cannot be due to radial pulsation. However, observations at high resolution must be made to search for the instability predicted by Maeder (1985) and Noels and Gabriel (1984), especially because of its intimate coupling to high mass-loss rates. The assumption of pure Rossby waves of low nonradial order, m, applied to stars with the shortest observed periods (0.5 d) may require rotation near breakup velocity. If this induces variations of the wind density, one might expect larger polarization variations than observations indicate thus far (Moffat et al., 1987b). Finally, periods of g^+ modes are given by Kirbiyik et al. (1984) for different stellar models; few hours to about 2 days (i.e., they span the observed range). Because unstable modes have not yet been found, more specific predictions are not possible. (However, Noels and Scuflaire (1986) are cautiously optimistic for low-degree, ℓ, modes in shell-burning models; unfortunately, the unstable phase found lasts only a few thousand years.) It would be all the more important to further test Vreux's (1985) suggestion that the intrinsic periods of Wolf-Rayet stars obey certain selection rules.

The most evident observational effect that NRP's have on the winds of O stars are the V/R variations of their emission lines (Section VIII.B.2.b). V/R variations can also occur in close binaries and are therefore not a suitable criterion for distinguishing between the two models. In any case, however, one would also expect to find them in Wolf-Rayet stars. However, inspection of emission-line profile series of the stars discussed by Vreux (1985) shows that profile variations are generally difficult to detect with photographic plates. Only HD 50896 (= WR6) shows some V/R-like variations (Ebbets, 1979b; Firmani et al., 1980). The line-profile variations reported in the two papers are in good qualitative accord and can be phased with the 3.76 d period derived and attributed to a binary with compact component by Firmani et al. Ebbet's high S/N data further show that different groups of lines behave very differently. Both data sets cannot clearly rule out periods shorter than 3.76 d, and neither can McLean's (1980) polarimetry, which yielded a roughly sinusoidal double wave when folded with the 3.76 d period. However, significant *stochastic* variations were also seen. In his discussion, McLean mentions NRP's, but does not consider any effects the pulsations may have on the geometry of the wind and therefore finds a binary model the most convincing. The relatively large amplitude of 0.3 percent may in fact tend to support this conclusion, but for that matter, it would be particularly desirable to observe HD 50896 again. Willis et al. (1986) comment on the similarity in the UV of this Wolf-Rayet star to massive X-ray binaries, but Willis et al. (1985) find rapid (~ hours) variations in the UV spectrum which are inconsistent with 3.76 d as the primary time scale.

If NRP's cause small discrete amounts of mass to be lost from the star, either periodically at every passage by a wave of a given point on the stellar surface or due to the interaction between several modes, they may also lead to detectable features in the profiles of wind lines (cf. Section VIII.B.3). The former case was recently modeled by Scuflaire and Vreux (1987) for a sectorial quadrupole mode ($\ell = 2$, $m = -2$). The effect of the resulting two-armed spiral on P Cygni profiles of resonance lines is a series of blueward propagating narrow components similar to the results of Prinja and Howarth (1985). Such variations have been observed in several He I absorption lines of HD 151932 (= WR78; WN7) by Seggewiss (1977), who has interpreted his results along the same lines but without referring to NRP's. At intervals of approximately 3.5 d, a new absorption component developed and was observed to accelerate at a constant rate of about -110 km/s/d from -700 to -1100 km/s. In He I λ3889, there was also a stationary component at -1270 km/s. Six years later, this behavior was no longer detectable (Seggewiss and Moffat, 1979), and no periodicity was seen that could be attributed to a companion. Furthermore, it is interesting to note that Lamers et al. (1985) also interpret the UV variability observed in P Cygni as a regular ejection of shells, although at intervals one hundred times longer.

Seggewiss' (1977) observations thus appear to be an encouraging result in the search for NRP's in Wolf-Rayet stars. However, the implied velocity law differs substantially from stellar-wind theories so that the velocity difference with respect to the ambient medium would be very large. As yet, there are no detailed calculations of the acceleration of relatively dense gas clouds in fast radiatively driven winds. This acceleration problem would be different if some matter initially corotated with a hypothetical companion in its Roche lobe and then got lost from the system through the outer Langrangian point so that the radial acceleration would set in at larger distances from the Wolf-Rayet star.

If Wolf-Rayet stars are nonradial pulsators and if the event-like long-term variations described in Section VIII.C.1 are a single-star phenomenon, it would be tempting to compare the latter to the *shell* ejections of Oe/Be stars, which: (1) also lead to nonspherically symmetric structures, (2) repeat in some stars quasi-periodically on time scales of years to decades, and (3) show a correlation with variations of the stellar NRP amplitude (cf. Section VIII.B.2.b).

D. Conclusions

Through the extrapolation of the early results summarized previously, evidence of NRP's in O stars should soon become ample. Observationally, the classification of the pulsation modes will form one of the most important tasks for the next few years, the main problem being that, because there is probably not one dominating restoring force, even the angular part of the eigenfunctions is not known. The empirical analysis of the atmospheric velocity fields and the continued theoretical search for a driving mechanism could be mutually beneficial. A thorough understanding of the pulsations will give a much firmer handle on the effects the pulsations have on the atmosphere, the structure of the wind, and the mass-loss rates. That such effects exist is qualitatively rather certain from the existing observations. The quantitative analysis still remains to be done. The variability that now appears to be the least clearly related to NRP's consists of the cyclic changes in the Hα emission strength of relatively narrow-lined O supergiants. This could be due to different stellar variabilities, but may also reflect the effect of different rotation rates. Nonaxisymmetric NRP's can provide a way to tap the huge reservoir of rotational energy and may therefore renew the interest in the rotation of early-type stars.

In contrast to the O stars, my evaluation of the current evidence of NRP's in Wolf-Rayet stars must appear to be less confident, and a further difficulty with the interpretation of short-term variations is that one first must ensure that they are caused by the pulsations of an O-type companion as in γ^2 Vel. However, the problems faced by pure binary models when confronted with some of the observed variations are not less severe—very short periods, low X-ray flux, time varying RV amplitudes, different RV amplitudes of different lines, doubtful long-term phase coherence, the possibility of multiperiodicity (Vreux, 1985, 1987; Smith et al., 1985), and uncorrelated spectral variations in the UV and the visible, to mention only the most prominent ones. Some of them show rather clearly that Wolf-Rayet stars are intrinsic variables, and for stable periods of the order of one-half a day, NRP's are the most plausible working hypothesis. Even considering that theoretical calculations have not yet detected a driving mechanism, the assumption of NRP's does not appear too contrived because: (1) NRP's are seen in several Of stars, (2) transition-type WNL stars are probably an advanced evolutionary stage of Of stars, and (3) the intrinsic photometric and polarimetric noise appears to peak at WN8. Because some WNL stars show intrinsic absorption lines at velocities comparable to the Hα absorption component of ζ Pup, similar variations as in this Of star may be detectable at high S/N and time resolution. A positive result should be the most convincing evidence of NRP's in Wolf-Rayet stars that one can currently think of.

E. Acknowledgments

I am indebted to Dr. Jean-Marie Vreux, who put the viewgraphs used for the presentation of his 1986 paper at my disposal. They were of great value in preparing the Wolf-Rayet star part of this work. I thank Drs. Alex Fullerton, Doug Gies, Tom Bolton, Don Penrod, Raman Prinja, and Ian Howarth for making available observations and results in advance of publication. Drs. Marc Azzopardi, Jacques Breysacher, Alex Fullerton, Leon Lucy, Raman Prinja, Myron Smith, Otmar Stahl, and Jean-Marie Vreux kindly read the manuscript. I am grateful for the comments provided, which resulted in a number of improvements. I also benefited from reading the manuscripts of the respective contributions to this volume sent to me by Drs. Peter Conti and Huib Henrichs.

4

STELLAR WINDS

I. INTRODUCTION

All O and Wolf-Rayet stars have highly ionized stellar winds, as evidenced by the ubiquitous appearance of P Cygni profiles of resonant lines of C IV, N V, and Si IV as observed with the Copernicus and IUE satellites (e.g., Figures 2-1 and 2-2). The maximum outflow velocities obtained from the violet absorption edges of the profiles are typically a few times the inferred escape velocities; thus the material is being lost from the star. This velocity is commonly considered the "terminal" velocity of the wind. The resonant lines of these ions are detected for a few (and up to some tens of) stellar radii away from the stellar "surface." However, the material in the wind extends considerably farther outwards. In some O stars, the free-free emission from the electrons in the winds at some few hundred stellar radii is detected by the radio continuum emission at 6 cm. Beyond this point, the stellar wind continues to flow but is not directly detected. At a distance of the order of parsecs, the boundary of the flowing material with the preexisting interstellar medium (ISM) is observed in the form of emission lines in such features as H, O II, O III, etc., which can sometimes be detected with direct imaging.

The previous sections discussed the observations and modeling that describe the underlying star and the stellar photosphere. This section considers the processes and data concerning the stellar wind itself. In Chapter 5, I will briefly discuss the issues concerning the boundary of the wind with the ISM.

Nearly all investigators agree that the wind flow is accelerated by radiative forces, which are very prominent in hot and luminous stars. The substantial momentum associated with the absorption of radiation by some ions is distributed to the atoms by ion/proton collisions, and a mass loss is inevitable. The "theory" was formulated by Castor et al. (1975a) in a seminal paper, which followed on the initial arguments of Lucy and Solomon (1970). The initial Castor et al. theory assumes a steady state for the wind. It is clear that the observed winds are not steady, and variability in O and Wolf-Rayet stars is found with various time scales (Section IV). Overall, it appears to me that the amplitude of the variability is such that it can be treated as a perturbation on a steady flow. In other words, O and Wolf-Rayet stars have an underlying steady radiatively driven wind in which variability is superimposed. This is to be contrasted with Oe and Be stars for which substantial time variations are found in the wind and material in the near vicinity of the star.

Considerable controversy still exists over whether or not the flow is "initiated" by the radiative forces (Abbott, 1982a) or by "subphotospheric" nonthermal processes not yet completely specified (e.g., Thomas, 1983). On one hand, a steady unvarying wind appears capable of being produced by radiative forces only (a sufficient but perhaps not a necessary condition); on the other hand, a specific "sub-photospheric" process such as nonradial pulsation (see Section IV) may be important in O and Wolf-Rayet stars (a necessary condition but perhaps not a sufficient one). This is an important issue that is not yet resolvable but may be more amenable to solution with considerable additional data, better physical insight, and improved modeling.

Let me briefly outline the radiatively driven wind theory. Originally, Abbott (1986) had written a review of this work for this NASA volume, based on remarks he gave at the *Third Trièste Workshop* in August 1984 (Abbott, 1985). Since that paper has been published in the proceedings of the workshop, an editorial decision was made to exclude its replication here. The interested reader may find Abbott's contribution in that reference, which discusses in more detail the current status of the radiatively driven wind models and the comparison to the extant observational data.*

The radiative acceleration is due to both the continuum and the line forces. The former, coming almost entirely from electron scattering in O and Wolf-Rayet stars, is itself substantial, and in these stars, the acceleration is 0.1 to 0.5 times the inward gravitational term. The value of the line acceleration and the resultant forces was initially approximated by Castor et al. (1975a). An accurate calculation of tens of thousands of lines, from an enormous accumulation of atomic data, was given by Abbott (1982a). In that paper, he gives the numerical value of the line acceleration for a large grid of models. According to Abbott, stars with effective temperatures above 10000 K have sufficient continuum and line forces to lead to winds. In cooler stars, the observed winds must clearly be the result of other physical processes (e.g., the solar case).

*Later in this chapter, Kudritzki et al. discuss radiatively driven stellar winds in more detail (Section IV).

The simplest formulation of the radiatively driven wind theory is to consider only the pressure terms for continuum and lines (the radiation), the gas pressure, and the gravity. The flow is assumed to be steady, laminar, and spherically symmetric. The several equations governing this case were given by Castor et al. (1975a), who then simplified them to two equations, one fixing the mass loss and the other the acceleration of the flow. In this idealized situation, a unique solution can be found, given the boundary conditions. More details of the process can be found in Abbott's (1986) paper and later in this chapter.

Real stellar winds are not always steady, laminar, or spherically symmetric. Various improvements to the basic Castor et al. (1975a) theory have been given by a number of authors. Abbott (1982a) gave a more realistic calculation of the line-radiation pressure. The assumption that the radiation was a point source was removed by considering a correct finite angular origin by Weber (1981), Friend and Castor (1983), and Pauldrach et al. (1986). This made a fairly dramatic change in the wind solution from the Castor et al. approach. In particular, the predicted terminal velocities turned out to be three times the escape velocity, as observed, rather than the smaller values predicted by Castor et al. Castor and Weber (private communication, 1984) dropped the Sobolev approximation in the radiation-transfer equations and found little difference in the wind solutions, except at subsonic velocities. Castor et al. considered only single scattering of line photons; multiscattering has been discussed by Panagia and Macchetto (1982b), who found that it may be important in stars with strong stellar winds, such as Wolf-Rayet stars. The effects of stellar rotation and a stellar magnetic field have been treated by Friend and MacGregor (1984). Both of these effects increase the mass-loss rates, but unless the rotations are near critical or the magnetic fields substantial, the resultant wind models are not much different from standard ones.

The present agreement between the improved radiatively driven wind models and some of the data is quite good. In particular, the prediction that, over a few orders of magnitude, the mass-loss rate should scale as the luminosity to a power (less than 2) appears to be verified (Section II, Figure 4-2, Abbott (1986), and Pauldrach et al. (1986)). Similarly, the prediction of the terminal velocity seems to be in accord with observations. In other words, the overall flow parameters appear to be predictable quantitatively (see Section IV later in this chapter).

Although the mass-loss rates scale with the luminosity as predicted by the radiatively driven wind theory, a certain amount of scatter occurs in the diagram. Part of this is attributable to observational error, partially in the luminosity and otherwise in the mass-loss rates themselves. The former is derived from apparent magnitudes, bolometric correction, and distances from cluster membership. Uncertainties of 0.2 in log L accumulate from these factors. The mass-loss rates probably have uncertainties of a factor of 2 given the inadequacies of the methods. These identifiable uncertainties are probably insufficient to explain the dispersion in the mass-loss rate/luminosity diagram (see Figure 4-2 in Section II), and a real dispersion probably exists. Later in this chapter, Kudritzki et al. give a plausible suggestion that the scatter is related to stellar evolution (Section IV).

One essential parameter of stellar winds which has not yet been well addressed by the Castor et al. (1975a) theory is the ionization state of the wind (e.g., Cassinelli, 1985). The ubiquitous appearance of the resonant N V and O VI ions in hot stars suggests that there is additional nonradiative energy input to the stellar winds. This is termed "superionization." The underlying photospheres of early B and late O stars have insufficient radiation to produce N V; similarly, even the hottest O stars should not show O VI lines. Anderson (1985) has begun a major effort to calculate line-blanketing without the local thermodynamic equilibrium (LTE) assumption. Using only the addition of carbon to the hydrogen and helium mixture adopted by Auer and Mihalas (1972), he finds an altered ionization balance in the wind, compared to LTE. While this non-LTE calculation begins to explain the presence of C IV in the early B-type stars, it is apparently not sufficient to produce the O VI. However, Pauldrach (1986) has recently given an improved non-LTE calculation for multilevel ions in a model for ζ Pup. He is apparently able to produce O VI lines without the necessity for superionization, although this result is currently controversial.

Another more common proposed solution to the superionization problem is to argue that the X rays present in hot-star winds (e.g., Harnden et al., 1979) produce these anomalous lines by Auer ionization processes (e.g., Cassinelli et al., 1978; Olson and Castor, 1981). The question then becomes: where are the X rays produced?

The X rays appear to be present in all hot luminous stars with luminosities of 10^{-7} of the bolometric value. There is a dispersion in the X-ray luminosity which amounts to a factor of ten from the mean value for OB stars, with a few exceptional larger values. The issue is thus one of producing the X rays in the stellar winds. Waldron (1984) constructed a model with a thin corona at the base of the wind, following previous suggestions of Cassinelli and Hearn. This potentially can explain at least some of the observations but has always appeared to be a little ad hoc.

That radiative flow models are potentially linearly unstable has been recognized (Lucy and Solomon, 1970; MacGregor et al., 1979). Lucy and White (1980) suggested that a nonstationary flow with two components to the wind might be appropriate; the resultant shocks would produce X rays. (More details are found in Lucy (1982a, 1982b).) Owocki and Rybicki (1984, 1985) have listed arguments for basic instabilities in line-driven stellar winds (see also Lucy, 1984). Krolick and Raymond (1985) have also addressed these issues. These authors are able to show that O VI lines can be produced by shocked stellar-wind material. My feeling is that this entire field of theoretical work is evolving rapidly and may prove to be critical to understanding of the ionization balance and X-ray production in the hot-star winds. We are temporarily stymied in acquiring additional spectroscopic data on O VI since that wavelength is unavailable to the IUE satellite (and the Space Telescope, unfortunately). More X-ray data will require additional satellites. The basic constraints on these two parameters are not well elucidated for OB stars or for Wolf-Rayet objects. This topic will undoubtedly undergo considerable revision in the near future, with additional physical arguments and with better data.

One other aspect of stellar winds has not been adequately addressed by the radiatively driven models: variability. It is possible that subphotospheric processes (Thomas, 1983) originate in the wind and so modify its character. One such process might be pulsations, both radial and nonradial, the latter of which have been demonstrated to be present in at least some O stars (Vogt and Penrod, 1983; Baade, Chapter 3 of this volume) and possibly in Wolf-Rayet objects (Vreux, 1985). The energy might be sufficient to modify the wind. Questions addressing whether the winds are ubiquitous in these objects, and the values of the energy input, need to be settled by more data than now exist. I believe that these are probably important in OB and Wolf-Rayet stars, but only to perturb the overall wind pattern, and to be the cause of the variability. In Oe and Be stars, I believe that these nonradiative processes dominate the resultant physical situation.

The Sections II and III of this chapter present the basic wind data known for the O-type stars and the Wolf-Rayet objects. Section IV contains a review of the stellar-wind theory by Kudritzki et al. Section V presents an overview of the situation as regards variability, and addresses the appearance of the narrow components in the wind. These data may be coupled to some common physical effect which is not yet well included in the models.

II. MASS LOSS FROM O STARS*

That some O stars are losing mass via stellar winds has been recognized since early rocket flights discovered P Cygni profiles in the luminous Orion supergiant stars with absorption component velocity shifts greater than the stellar escape velocity (Morton, 1967). However, it was the advent of ultraviolet (UV) spectroscopy which established that practically all O stars are losing mass. The Copernicus satellite led to the first quantitative measurement of mass loss, and subsequent observations with the IUE satellite have extended these results. At the other end of the spectrum, radio observations with the very large array (VLA) have been used to make independent determinations of mass loss from the most luminous O stars. There have also been some determinations of mass loss from Hα modeling. Although early work at infrared wavelengths seemed to be a promising method of deriving mass loss, this has not turned out to be the case, although other important information about the wind velocity law is contained within the infrared (IR) observations. As each method is model-dependent to some extent, a very important test of the underlying assumptions has been the comparison of rates for the same star obtained by different methods.

Earlier comparisons of mass-loss rates derived by different methods did not always agree very well, leading to speculation that the theory of stellar winds was inadequate or that problems were present in the methods. For example, Abbott et al. (1981) found that their radio mass-loss rate did not agree very well with Hα determinations. However, when it was realized that some O stars were nonthermal radio sources, the worst discrepancies between the radio mass-loss rates and the Hα mass-loss rates were explained. For example, this problem arose in the comparison of radio and UV rates

*C. D. Garmany, Joint Institute for Laboratory Astrophysics, University of Colorado and the National Bureau of Standards, Boulder, Colorado, 80309-0440.

in 9 Sgr, which was later recognized as a nonthermal source.

The following sections discuss the primary source for mass-loss determinations by different methods. At present, mass-loss values for about 60 galactic OB stars are known. The rates of some of the stars were determined by multiple methods, which permit a comparison of the accuracy of the methods.

Table 4-1 lists the parameters for the stars with mass-loss rates, as well as the rates, terminal velocities, and other wind parameters. The spectral types (column 2) are from Walborn (1972, 1982c) when available. In the case of cluster members, the bolometric magnitudes are derived from absolute visual magnitudes based on cluster distance from Humphreys (1978). An "f" following the bolometric magnitude indicates a field star, in which case the absolute magnitude is based on the spectral-type calibration discussed in Chapter 2 of this volume. The bolometric corrections and the temperatures are based on the discussion in Chapter 3. Note that these are not necessarily the same temperatures and luminosities used in the original derivation of the various mass-loss rates. In general, because the temperatures here are lower, especially for the hottest stars, the bolometric corrections are also lower. In most cases, this change would not have a major effect on the value of the derived mass loss. Columns 5, 6, and 7 include wind parameters. Column 5 gives the terminal velocity of the wind as derived from UV resonance-line P Cygni profiles observed with either IUE or Copernicus. Some stars with measured terminal velocities or other wind parameters, but no mass-loss rates, are included to give a better impression of the dispersion of terminal velocities. Column 6 lists T, a parameter defined by Castor and Lamers (1979) which is a measure of the column density of the wind. In this case, it refers to the strength of the resonance line of C IV 1548,1552. In general, a value of T = 20 indicates that the line is completely saturated. Column 7 is the classification of the strength of the luminosity-sensitive Si IV 1393,1402 lines used by Walborn and Panek (1984a). A 3 refers to a P Cygni profile, and a 0 refers to a photospheric line in their notation.

Columns 8, 9, and 10 of Table 4-1 give mass-loss rates, in logarithmic units of solar masses per year, from various sources. Column 8 includes UV determinations, discussed in Section II.A. If two values are given, the top one is from either Olson and Castor (1981), Garmany et al. (1981), or Garmany and Conti (1984), and the bottom one is from Gathier et al. (1981). Column 9 shows radio rates, discussed in Section B, and column 10 gives rates derived from Hα measurements, discussed further in Section C. These rates include only primary sources (i.e., they do not include later compilations which have adjusted or scaled mass-loss rates based on these rates). Care must be taken in intercomparing methods to use only primary sources, or one can fall into the trap of comparing the same mass-loss determination with itself!

A. Ultraviolet Mass-Loss Determinations

With the advent of UV astronomy, it became clear that all O stars exhibit mass loss. P Cygni profiles, observed in only a very few cases in the visible part of the spectrum, are seen in almost every O star observed thus far. The shape and strength of a P Cygni profile is determined by the mass-loss rate and the velocity law and ionization of the material as a function of radius. The strength of the blue-shifted absorption component of the profile is a measure of the column density of the ion and is not very dependent on the velocity law. The ratio of the emission component to the absorption component depends on the shape of the velocity law because UV resonance lines are largely scattering lines. Therefore, this ratio depends on the distance of the absorbing material from the star. Mass-loss rates can be computed from P

Table 4-1
Mass-Loss Parameters in O-Type Stars

Star	Spectral Type	$-M_{bol}$	Temp.	$V_\infty/10^3$	T C IV	Walborn Si IV	UV	log $\dot{M}$ Radio	Other
93205	O3 V	9.7	48.0	2.85					
93129A	O3 If	10.5	43.0	3.9:					
93250	O3 V((f))	10.6	48.0	3.50	5	0	-5.9		
303308	O3 V((f))	9.6	48.0	3.40	≥20	0	-5.6		
46223	O4 V((f))	9.4	45.0	3.10	≥20	0	-5.6		
9 Sgr = 164794	O4 V((f))	10.2	45.0	3.50	≥20	0	-5.3	N.T.	
							(-5.5)		
242908	O4 V(n)	9.4	45.0	3.2	8	0	-6.1		
ζ Pup, 66811	O4f	9.8	41.0	2.66			-5.45	-5.42	-5.24
							-5.53		
Cyg OB2 #7	O3f	10.0	43.0	(3.8)				-4.72	
HD 15570	O4f	10.5	41.0	2.7				-5.00	
15629	O5 V((f))	9.4	43.0	3.20	≥20	0	-5.6		
46150	O5 V((f))	9.1	43.0	3.20	8	0	-6.1		
93204	O5 V((f))	8.8	43.0	3.20	≥20	0	-5.7		
15558	O5 III(f)	10.2	41.0	3.0		0			
14434	O5 V	8.6	43.0	2.2	>20	0	>-6.2		
Cyg OB2 #9	O5f	10.9	39.0	(2.65)				-4.72	
101190	O6 V((f))	9.4	41.0	3.10	8	0–1	-6.0		
101298	O6 V((f))	9.0	41.0	3.00	≥20	1	-5.8		
-59°2600	O6 V((f))	8.9	41.0	3.30	10:	1	-6.0		
152233	O6(f)	9.8	39.0	3.20	≥20	1	-5.4		
λ Cep, 210839	O6ef	9.7	36.5	2.50	≥20	3	-5.4		
Cyg OB2 #5	O6f + O7f	(1.8)						-4.55	
12993	O6.5 V	8.1	40.0	2.5	4	0	-6.6		
42088	O6.5 V	8.2	40.0	2.60	4	0	-6.9		
54662	O6.5 V	9.1	40.0	2.50	3	0	-6.7		
101436	O6.5 V	9.1	40.0	3.10	≥20	0–1	-5.9		
206267	O6.5 V((f))	9.2	40.0	3.10					
0163758	O6.5 Iaf	9.9f	35.5	2.6				-5.2	

Table 4-1 (Continued)

Star	Spectral Type	$-M_{bol}$	Temp.	$V_\infty/10^3$	T C IV	Walborn Si IV	UV	log $\dot{M}$ Radio	Other
199579	O6 V((f))							-5.9	
152723	O6.5 III(f)	9.3	38.0	3.64	6	1	-6.1		
190864	O6.5 III(f)	9.1	38.0	2.95	≥20	1	-5.7		
-59°2603	O7 V((f))	8.3	39.0	2.4:	1		-7.2		
36879	O7 V(n)			1.5:		pec, var			
15 Mon, 47839	O7 V((f))	8.4	39.0	2.2:		0	-6.82 -6.27		≤-6.22
48099	O7 V	8.9	39.0	3.50	6	0	-6.2		
152623	O7 V(n)((f))	9.6	39.0	3.25	4	1–0	-6.3		
93222	O7 III((f))	8.4	37.0	2.80	10	0–1	-6.3		
167659	O7 II(f)	9.3	34.5	2.60	≥20	1	-5.5		
151515	O7 II(f)	9.1	34.5	2.70	>20	2	≥-5.8		
152248	O7 I	9.9	34.5	3.08	≥20	3			SB2
35619	O7 V	9.2	39.0	2.32	3		-7.3		
68 Cyg, 203064	O7.5 IIIn((f))	8.8	35.0	2.70					
152590	O7.5 V	8.0	37.5	2.0	3	0	-7.4		SB
53975	O7.5 V	8.3	37.5	2.0		0			
ξ Per, 24912	O7.5 III((f))n	8.4	35.0	2.5		2	-5.80		
9 Sge, 188001	O7.5 Iaf	10.1f	33.0	2.3			-5.2		
14633	ON8 V	8.0f	36.5	2.4:		0			
46966	O8 V	8.2	36.5	2.3		0			
48279	O8 V	8.0f	36.5	2.3:		0			
101413	O8 V	8.1	36.5	2.85	3	0	-6.9		
λ Ori A, 36861	O8 III((f))	8.4	34.0	2.30		1	-6.70	-6.10	<-6.22
175754	O8 II((f))	8.8f	32.0	2.30		3			
46056	O8 V(e)	7.6	36.5	1.60	2		-7.6		
151804	O8 If	10.3	32.0	2.0	>20	3	-5.1	-5.03	
152408	O8 If	10.0	32.0	1.8	>20	3	-5.0	-4.74	
46149	O8.5 V	8.1	35.0	1.70	2	0	-7.7		
193322	O9 V:((n))	7.5f	34.0	1.8:					
τ CMa, 57061	O9 II	9.9	29.5	2.3			-5.99		

Table 4-1 (Continued)

Star	Spectral Type	$-M_{bol}$	Temp.	$V_\infty/10^3$	T C IV	Walborn Si IV	UV	log $\dot{M}$ Radio	Other
10 Lac, 214680	O9 V	7.6	34.0			0	−6.59		
57682	O9 IV	7.4f	34.0			0			
ι Ori, 37043	O9 III	8.9	32.0	2.7:		1	−6.52		<−6.05
							−5.94		
152246	O9 III	8.6	32.0	2.3	6		−6.9		
210809	O9 Iab	9.1	29.5	2.1					
207198	O9 Ib-II	8.4	2.95	2.3					
AE Aur	O9.5 V			0.8:		0			
μ Col, 38666	O9.5 V	6.5	33.0	1.2:		0	−8.22		
							−7.17		
δ Ori A, 36486	O9.5 II	9.3	28.0	2.3		2	−6.30	−6.03	
							−6.10		
ξ Ori A, 37742	O9.7 Ib	9.7	28.0	2.2		3	−6.30	−5.64	−5.64
							−5.80		
47432	O9.7 Ib	8.7f	28.0	2.2		3			
152314	O9.5 III	7.8	30.5	2.62	$\geq$20	3	$\geq$−6.0		
152424	O9.5 I	9.8	28.0	2.24	$\geq$20	3	−6.1		
152247	O9.5 II	8.5	28.0	2.6	$\geq$20		−6.1		
13745	O9.5 II	8.0	28.0	2.4	$\geq$20	2	−6.1		
α Cam, 30614	O9.5 Ia	8.7	28.0	1.75			−5.29		−5.45
218915	O9.5 Iab	8.7f	28.0	2.4					
154368	O9.5 Iab	8.7f	28.0	2.3					
188209	O9.5 Iab	8.7f	28.0	2.1					
152249	OC9.5 Iab	9.0	28.0	2.2			−6.2		
209975	O9.5 Ib	8.4	28.0	2.3					
14947	O5f	9.7	39.0	2.7	$\geq$20		−5.4		
166734	O7.5f + O9 I	9.5	33.0	2.6				−4.68	
149757	O9.5 V						−6.48		

Cygni profiles because the mass-loss rate depends on the ion column density, the ionization fraction and abundance, and assorted atomic parameters. The other required parameters include the wind terminal velocity, which is measured directly from the blue edge of the absorption P Cygni profile, and the stellar radius, which is derived from the luminosity of the star.

Castor and Lamers (1979) discuss the derivation of mass-loss rates from the equation of continuity:

$$\frac{dm}{dt} = 4\pi\, R_*^2\, \frac{dx}{dt}\, \delta_{gas} \quad , \qquad (4\text{-}1)$$

and the optical depth of a line in an expanding envelope:

$$\tau(r) = \frac{\pi e^2}{mc}\, g\lambda n_i \left(\frac{dr}{dv}\right) \quad . \qquad (4\text{-}2)$$

The task of actually deriving mass-loss rates has been greatly simplified by the series of line-profile models parameterized by Castor and Lamers (1979) and further modifications by Olson (1982) to include doublet profiles. The primary sources of mass-loss rates determined from fitting UV resonance P Cygni profiles are Gaither et al. (1981), Olson and Castor (1981), Garmany et al. (1981), and Garmany and Conti (1984). These papers include a total of 67 O stars. The particular resonance lines used are not the same in each of these papers: the first two include Copernicus spectra and therefore extend farther into the UV than the latter two papers, which are based on spectra obtained with the IUE satellite.

The method used by Gaither et al. (1981) was an empirical one: Copernicus profiles for about five ions were fit to theoretical profiles of Castor and Lamers (1979) to derive column densities, and the column densities were calibrated by means of IR or radio rates from ten of the stars. At present, there is great controversy about IR rates, although the radio rates used in their calibration are reliable according to Abbott (1985).

The method used by Olson and Castor (1981), Garmany et al. (1981), and Garmany and Conti (1984) is rather more model-dependent than any other method of obtaining mass-loss rates. Although the line-profile fitting, which is done with a maximum of two or three parameters, is model-independent, the computation of the ionization fraction for each ion is not. The rates derived by Olson and Castor (1981) were done by computing the ionization equilibrium of many elements and matching the line strengths of the UV resonance lines observed by the Copernicus and IUE satellites. The dominant stages of ionization in the winds of early-type stars are N IV, Si V, and C V (Olson and Castor, 1981), but these stages are not observed in the spectral region covered by IUE. This necessitates large correction factors when only IUE observations are available. It is necessary to observe two lines on opposite sides of the dominant ion, such as N V and C IV, in order to compute the correction. Because the largest systematic uncertainties are introduced in the ionization balance computations, it is very important to be able to compare the mass-loss rate from radio and UV for the same star.

B. Radio Mass-Loss Determinations

Radio observations of thermal emission from hot stars can yield both mass-loss rates and temperatures. The theory was developed by Panagia and Felli (1975) and by Wright and Barlow (1975). The stellar mass-loss rate is determined by the flux observed at radio wavelengths, the distance of the star, and the terminal velocity of the wind, generally measured from the short wavelength edge of a saturated P Cygni profile in the UV spectral region. A good review of the theory is given by Abbott (1985).

Implicit in the method are the assumptions that the wind outflow at the point where the free-free emission is detected is isotropic and at constant velocity and that the wind temperature and ionization are constant at large radius. Because all indications are that, at 6 cm, the radius of the radio photosphere exceeds that of the optical photosphere by a factor of 10 for O stars and by a factor of 1000 for Wolf-Rayet stars, these assumptions are well grounded. Other assumptions which were initially accepted without much question include the assumption that the radio emission is entirely thermal and the assumption that the star is not a multiple system. This latter assumption is based on the improbability of low-mass companions around O stars (Garmany et al., 1980). The former assumption has proved not to be the case in about one quarter of the OB stars.

Radio observations of O stars can be done only with the VLA and are therefore limited to northern hemisphere O stars with rates in excess of 10^{-6} M/yr. The primary data are contained in Abbott et al. (1980, 1981, 1984, 1985). A total of ten O-type stars have radio mass-loss rates that are considered to be due to either definite or probable thermal emission. Another five stars are nonthermal sources and therefore cannot be used to determine mass-loss rates.

C. Visible Mass-Loss Determinations

In the visible, the evidence for mass-loss comes mainly from the emission lines of hydrogen and helium. These lines are in emission only when the mass-loss rate is relatively large, greater than 10^{-6} M/yr^{-1}. For smaller rates, the only effect is a partial filling in by wind emission of the cores of photospheric lines, which is detectable as a weak or asymmetric profile. Klein and Castor (1978), using the Castor et al. (1975a) model of radiation-driven winds, derived expressions for mass loss as functions of the mass of the star and its Hα luminosity. Observations of the Hα equivalent width are then used to derive the mass-loss rate and compare it with the theoretical predictions which include a linear relation between the Hα luminosity and the mass loss. The same technique was used by Conti and Frost (1977), although different methods were used to derive the stellar masses. This method must assume the correct velocity law, which, according to recent IR work, varies from star to star. Recent detector technology now makes it possible to observe the weak emission wings of Hα, and profile fitting yields both the mass loss and the velocity law simultaneously. This method was used by Olson and Ebbets (1981) in a study that included six O stars.

Olson and Ebbets (1981) calculated theoretical Balmer line profiles in a spherically symmetric expanding wind using the Sobolev approximation. A simple correction for stellar rotation was included. The parameters that can be varied include mass loss, the velocity law, and the effective temperature. Knowledge of the wind's terminal velocity, from UV observations, is required for the adopted velocity law. Other physical parameters which are required include the star's effective temperature and radius. For the two stars in common with Klein and Castor (1978), the rates agree very well, suggesting that the simpler technique is adequate.

An interesting problem for rates derived was raised by Ebbets (1982), who studied the variability of the Hα profile and found that, for a large sample of stars, the Hα equivalent width is close to the predictions of Klein and Castor (1978), but that fluctuations in Hα lead to variations in the mass-loss rate from 5 percent in the course of days to 30 percent over weeks or months.

D. Infrared Mass-Loss Rates

Johnson (1967) was the first to suggest that the additional flux at 3.4 μm over the expected flux of a blackbody was due to a circumstellar shell. This is an important indication of an extended stellar envelope. Although early results of mass-loss determinations from the continuum free-free emission at IR wavelengths appeared promising, it now seems that such rates are highly uncertain. The IR excess measured for O stars comes from a region of the wind near enough to the star that knowledge of the velocity law must be in hand if a mass-loss rate is to be determined. Abbott et al. (1984), in a combined analysis of new IR data and their previous radio observations, found that the velocity law varies dramatically from star to star. The same conclusion was reached by Castor and Simon (1983), who found that, because the majority of 50 stars which they observed in the IR had winds that are optically thin in the near IR, these observations alone could not define the velocity law.

Because free-free opacity is proportional to λ, one can measure the wind at larger radii by measuring at longer wavelengths. However, there are practical problems with IR observations. O stars are weak sources in the IR, and the correction for interstellar extinction is often large even in the IR. A reddened spectrum can mimic the appearance of IR excess. IR absolute calibrations have about a 10-percent uncertainty, and the excess is measured by comparing with a model-dependent photospheric radiation field.

Abbott, et al. (1984) observed ten O stars in the *K, L, M, N*, and *Q* bands. All of the stars in their sample have radio mass-loss rates, and they do not attempt to derive mass-loss rates from the IR data. There is a trend for the velocity law to become more gradual as the temperature decreases, but for their sample of stars, this could also mean as the gravity decreases.

The conclusions of Persi et al. (1983) that their IR mass-loss rates differ from the radio rates and show no correlation with luminosity probably reflects the difficulty in using only IR rates for mass-loss determination.

E. Comparison of Mass-Loss Rates Obtained by Different Methods

In Table 4-1, there are 18 stars with multiple mass-loss-rate determinations. Figure 4-1 plots the different rates as a function of bolometric magnitude, and clear systematic differences can be seen for stars less luminous than M_{bol} = -10. Of the UV methods, the empirical method of Gathier et al. (1981) gives systematically larger rates by about 0.4 in the log than the method of Garmany et al. (1981). The difference cannot be settled by the radio rates because these are available only for the most luminous stars. The rates from Hα modeling are larger than those from both UV and radio for all luminosities. The systematic differences in Figure 4-1 make it clear that averaging multiple rates for the same star will not result in an improved mass-loss rate.

F. Dependence of Mass Loss and Luminosity

In the initial derivations, the theory of radiation-driven winds (Castor et al., 1975a) predicted that mass loss should depend on luminosity to the power 1, and early infrared mass-loss determinations appeared to confirm this relation (Barlow and Cohen, 1977). As more UV and radio rates became available, it became clear that the dependence of mass loss on luminosity was closer to the power 2. Both Abbott et al. (1981) and Garmany et al. (1981) found that mass-loss scales with luminosity to the power 1.6 or 1.7. The Gathier et al. (1981) data indicated a dependence on luminosity to the power 1.4. The possibility that other parameters were involved in the parameterization of mass loss has also been explored. Using rates derived from a variety of methods and corrected to a uniform set of estimated stellar parameters, Lamers (1981) derived a dependence of the form:

$$\log \dot{M} \propto 1.42 \log L_* + 0.61 \log R_* - 0.99 \log M_* \quad (4\text{-}3)$$

This can be compared with Abbott's (1982a) discussion of radiation-driven wind theory using a more realistic treatment of line acceleration. His relation takes the form:

$$log\ \dot{M} \propto 1.98 \log L - 1.03 \log [M(l\text{-}r)] + 0.94 \log Z - 0.02\ 4.4 \log T \quad (4\text{-}4)$$

Garmany and Conti (1984) explored empirical dependences and found that mass loss can be parameterized by luminosity and the ratio of the terminal velocity to the escape velocity, although the standard deviations of the uncertainties in the derived mass-loss rates do not

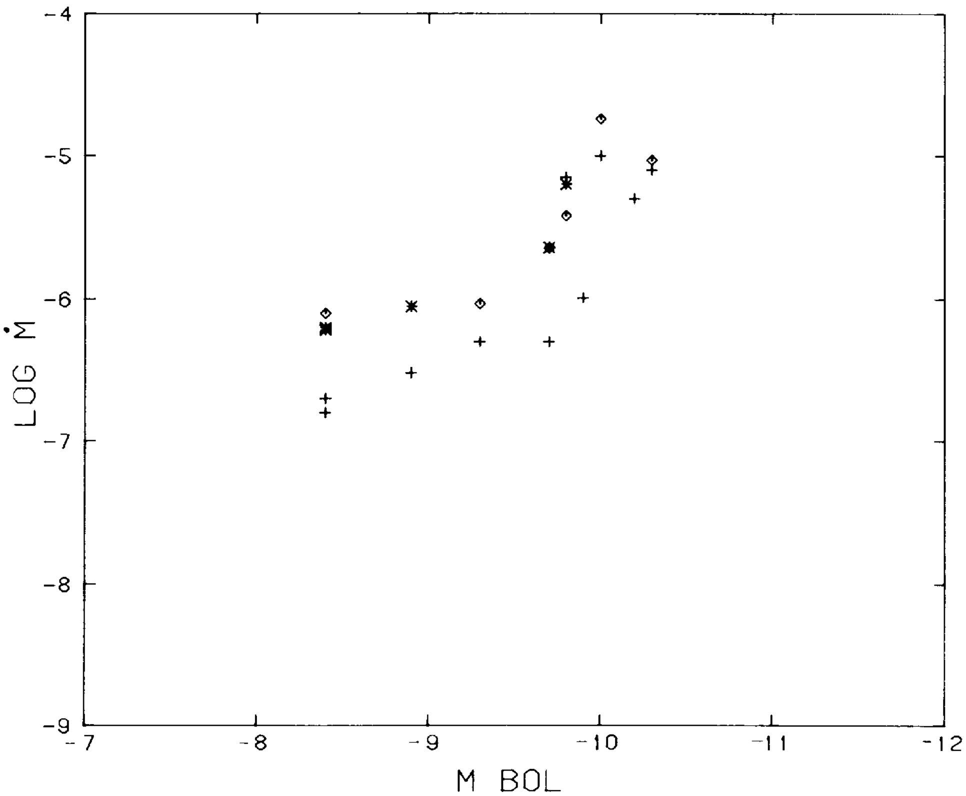


*Figure 4-1. A comparison of different methods of mass-loss determination for the same star as a function of bolometric magnitude: + denotes UV mass-loss determination, ◊ denotes radio determination, and * denotes visible determination.*

significantly recommend this additional parameterization over a straight dependence on luminosity.

All attempts to explore mass-loss parameterizations depend on estimated physical parameters for the stars. The most important one is the luminosity. For field stars, the uncertainty in the absolute magnitude calibration can translate into a factor of 1.6 uncertainty in the luminosity, so cluster members are much more reliable. Figure 4-2 shows mass loss versus bolometric magnitude for the data given in Table 4-1. The scatter in the relation is greater than one expects, either from the errors in mass-loss rate or errors in cluster distances, but without accurate knowledge of parameters such as temperature, bolometric corrections, mass, and metallicity, exploring their dependence on mass loss is very difficult.

III. MASS-LOSS RATES IN WOLF-RAYET STARS

Mass-loss rates in Wolf-Rayet stars can, in principle, be obtained from spectroscopic information such as that contained in the emission lines in the visible, or the numerous P Cygni lines there, and in the far-UV regions accessible to the IUE satellite. However, it is well recognized that this procedure is highly model-dependent (Section II) for O-type stars and even less trustworthy for these objects in which the complications of anomalous composition would also have to be dealt with. In Wolf-Rayet

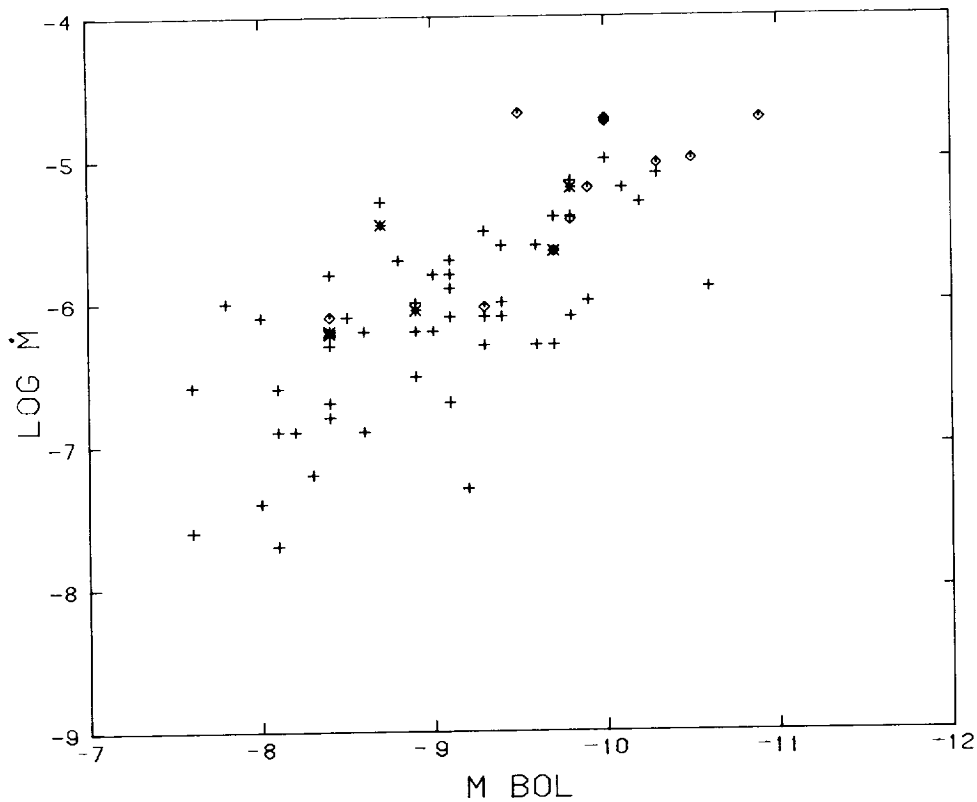


*Figure 4-2. Mass loss versus bolometric magnitude for all the data given in Table 4-1: + denotes UV mass-loss determination, ◊ denotes radio determination, and * denotes visible determination.*

stars, these considerations are paramount. For this reason, the mass-loss parameterizations have come almost entirely from the measurements of free-free emission in the IR or radio regions. As noted previously, the interpretation of the IR measurements is complicated by the necessity of adopting a velocity law before any conclusions can be drawn. For the radio data, the emission comes from far out in the wind where the terminal velocity is presumed to be reached and estimated from UV P Cygni profiles, and these ambiguities are not present.

Abbott et al. (1985, 1986) have completed a detailed study of detections of thermal free-free emission from the winds of Wolf-Rayet stars using the VLA. Their data are therefore limited to those Wolf-Rayet stars that are accessible to this northern hemisphere location. The observations are mainly at 6 cm wavelength, but some data at 2 cm are also available. The emission measure is proportional to the density in the wind, with practically no model dependence. The $\dot{M}$ can be found if the distance of the star is known (the uncertainty here scales as the 3/2 power) and if the terminal velocity of the wind can be obtained (typically with IUE satellite measurements). As for the O stars, a presumption is that the terminal velocity measured at a few tens of stellar radii is constant out to a few hundred radii where the free-free emission at these wavelengths is produced. In real stars, this assumption should be reasonably well satisfied.

Recently, it has come to be recognized that some of these hot luminous stars have a non-thermal component to the radio-emission measurements. Abbott et al. (1985, 1986) are able to show that for, most Wolf-Rayet stars, the nonthermal contribution is small because the wavelength dependence of the flux follows a power law with exponent -0.6. Abbott et al. list detections of seven Wolf-Rayet stars as definite thermal emitters and another 17 as probable thermal sources, thus giving 24 stars with well-determined mass-loss rates. They also compiled terminal velocities, and upper limits are available for nine additional stars. These data are listed in Table 4-2. Is there any dependence of $\dot{M}$ on the Wolf-Rayet subtypes? Figure 4-3, adapted from Abbott et al., suggests that there is not. The WN and WC subtypes appear to have similar $\dot{M}$, and there is no dependence on excitation class. There is a certain spread in the values of $\dot{M}$ at a given subtype, a point to which I will return shortly.

Abbott et al. (1986) have obtained $\dot{M}$ for five binary systems for which reasonable mass estimates are in hand. Figure 4-4, adapted from their paper, shows that a relation exists between the $\dot{M}$ and the stellar masses, with the mass-loss rates being larger in the larger mass stars.

Why does the $\dot{M}$ depend on the stellar mass? Abbott et al. suggest that this is because the more massive stars are the more luminous. As I have discussed previously, the bolometric luminosities of Wolf-Rayet stars are not well estimated. Abbott et al. were able to skirt this problem by calculating a mean M_v for each spectral subtype and a deviation from the mean for each Wolf-Rayet star with a measured $\dot{M}$. The correlation of this artificial luminosity with $\dot{M}$ is shown in Figure 4-5. The stars with the higher $\dot{M}$ tend to be brighter (in M_v within their subtypes). Figure 4-5 is currently the best indication that the mass-loss rate in Wolf-Rayet stars depends on the luminosity, which would be expected if the winds in Wolf-Rayet stars are radiatively driven, but it does not prove this connection. It is useful here to recall the recent theoretical model of Pauldrach et al. (1985), in which they are able to match the wind law of V444 Cyg found by Cherepashchuk et al. (1984)

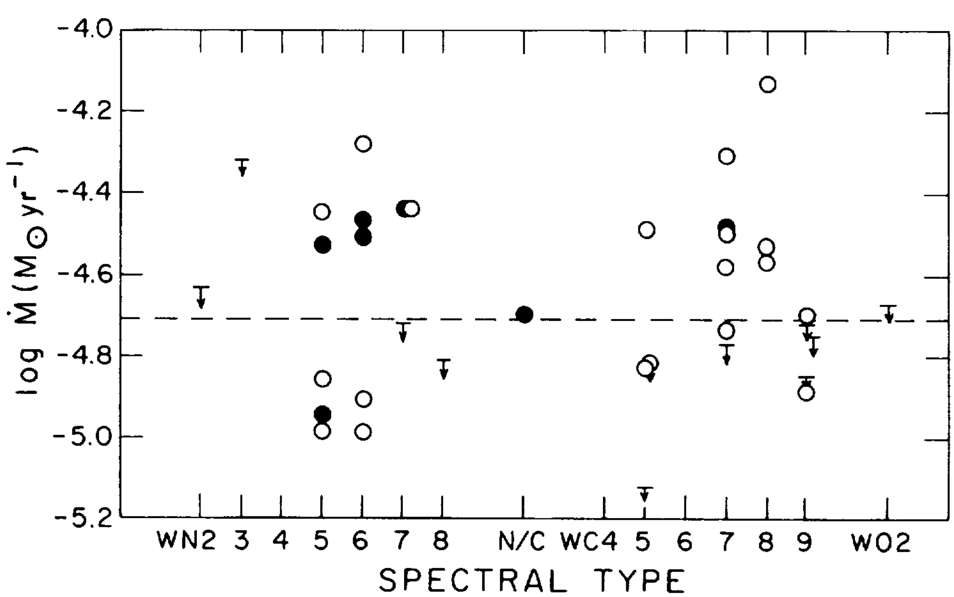


Figure 4-3. Mass-loss rates as a function of spectral type (from Abott et al., 1986): Filled circles denote definitive thermal wind sources and open circles denote probable thermal wind sources. There is no apparent correlation of $\dot{M}$ with subtype.

Table 4-2
Mass-Loss Parameters in Wolf-Rayet Stars

Star Name	Spectral Type	V_∞ km s^{-1}	log $\dot{M}$ $M_\odot$ yr^{-1}
Sand 5	WO 2	7400	<4.7
HD 165763	WC 5	3550	−4.8
HD 169010	WC 5	2600	<−4.8
HD 157502	WD 7 + ABS	3100	−4.5
HD 152270	WC 7 + 06	3300	−4.3
HD 156327	WC 7 + ABS	2400	−4.6
HD 156385	WC 7	3000	<−4.8
HD 192641	WC 7 + ABS	2700	−4.7
γ^2 VEL	WC 8 + 09 I	2000	−4.1
HD 168206	WC 8 + 08 V	2900	−4.5
HD 192103	WC 8	1900	−4.6
MR 66	WC 9	1300	−4.9
MR 74	WC 9	1300	<−4.8
HD 164270	WC 9	1400	$\lesssim$−4.7
MR 80	WC 9	1600	<−4.7
MR 90	WC 9	1300	<−4.8
HD 6327	WN 2	4200	<−4.6
HD 190918	WN 4.5 + 09 I	1750	⩽−5.0
HD 4004	WN 5	2850	−4.5
HD 50896	WN 5	2650	−7.5
HD 193077	WN 5 + ABS	1700	−4.9
HD 193576	WN 5 + 06	2500	−5.0
HD 165688	WN 6	3800	−4.3
MR 87	WN 6	1200	−5.0
MR 89	WN 7	2000	−4.4
LS 16	WN 8	1100	<−4.8
MR 111	WN/C	1500	−4.7

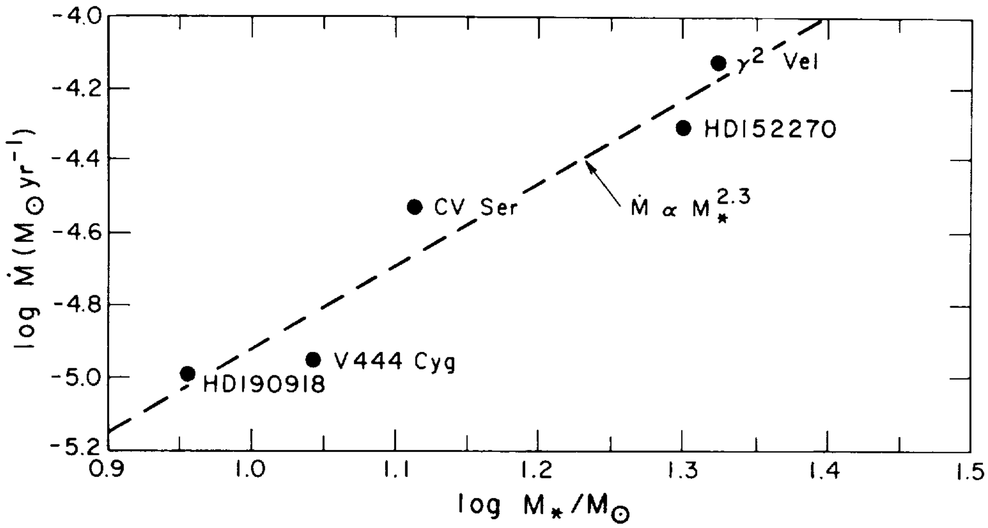


Figure 4-4. Mass-loss rates versus stellar mass for five Wolf-Rayet stars in which the latter quantity is determined from the solution of a double-lined spectroscopic orbit (from Abbott et al., 1986). The dashed line shows a tentative scaling.

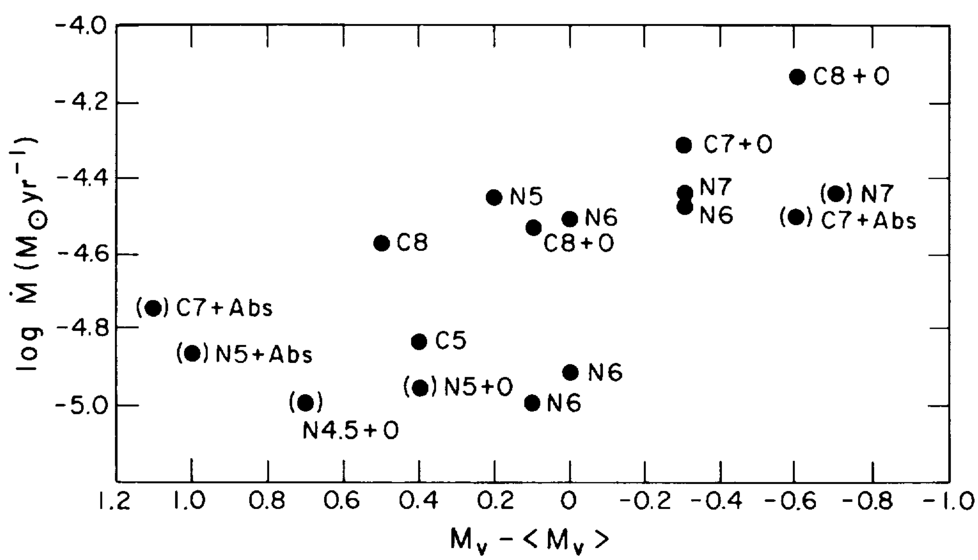


Figure 4-5. Mass-loss rates versus absolute visual brightness of Wolf-Rayet stars (from Abbott et al., 1986). The M_v - < M_v> quantity is the difference between the measured M_v of the star and the mean M_v of that subtype, proportional therefore to the luminosities. Parentheses denote less certain M_v values. A rough relationship between $\dot{M}$ and M_v - <M_v> is seen.

with a radiatively driven wind from a core of 90000 K effective temperature.

The substantial mass-loss rates found for Wolf-Rayet stars by Abbott and his associates are significant on the evolution time scale for these objects. These values will decrease by a considerable fraction the remaining mass in the helium-burning lifetime. As the star gets less massive, its luminosity may also decrease, and the mass-loss rate may taper off if $\dot{M}$ depends on L. The final state of a Wolf-Rayet star is by no means clear; whether they just "poop out" or violently explode is an interesting question.

Is it really possible that these large mass-loss rates are a result of the high effective temperatures, thus very large luminosities, and that the winds are driven by radiative forces? I believe that this is plausible, but will await further work. The wind law (i.e., the run of opacity and velocity with radius) in Wolf-Rayet stars appears to be rather "soft" as compared to OB stars. By this I mean that the wind velocity increases rather slowly with radius. What physically causes this difference as compared to the OB stars? I have no compelling suggestions except to note that the hydrogen/helium ratio, which is very low in most Wolf-Rayet stars as compared to the O types, may play a role in the structure of the wind, given the different opacity sources.

IV. RADIATION-DRIVEN WINDS OF HOT LUMINOUS STARS*

A. Introduction

Stellar winds are a ubiquitous feature of hot luminous stars. They dominate the ultraviolet spectra through the presence of broad velocity-displaced absorption lines and modify the photospheric energy distributions by additional free-free emission in the far-IR and radio domain. The evolution of massive stars is strongly affected by the severe mass loss connected with the stellar winds, which is reflected in the HR diagrams of massive stars in the Milky Way, as well as in the Magellanic Clouds and more distant galaxies. The stellar wind carries nuclear burned material back to the interstellar medium and provides an appreciable amount of energy that influences the energy balance and causes additional star formation. Moreover, as recently pointed out by Abbott and Hummer (1985), the stellar winds also modify the structure of the underlying hydrostatic photospheres of the stars by scattering back a large part of the continuous photospheric radiation. The effect of this "wind-blanketing" is that the spectral types of massive stars depend not only on effective temperature and gravity, but also on the density and extension of the surrounding stellar-wind envelope. In practice, this means that effective temperature, gravity, *and* mass-loss rate must be determined simultaneously when hot stars are analyzed.

Consequently, the development of a quantitative theory for the winds of hot stars is of crucial astrophysical importance. Until now, the most promising attempt was the theory of radiation-driven winds. It is the primary intention of this paper to demonstrate that this frequently denigrated theory is much better than its reputation. We will show that most of the basic observational properties of hot-star stellar winds, at least for massive luminous OB stars, can be explained by this self-consistent theory. Section B describes the framework of radiation-driven wind theory (including its very recent developments) and the basic observations that must be explained. In describing the theory, we have tried to find a simple way to explain how it works because we realize that most of the reluctance in accepting it comes from its formally complicated structure. Section C gives a comparison between the observations and the results of the improved theory. It includes the winds of massive O stars in the Magellanic Clouds and thus gives information about the metallicity-dependence of radiation-driven

*R. P. Kudritzki, A. Pauldrach, and J. Puls, Universitäts-Sternwarte München, Scheinerstr. 1, D-8000 München 80, Federal Republic of Germany.

winds. Section D discusses the reliability of the Sobolev approximation for the stellar-wind dynamics. Section E treats a longstanding problem: superionization and statistical equilibrium in stellar winds. It is shown that this problem can be solved in "cool winds" of O stars without any other source of ionization if a correct multilevel treatment of the statistical equilibrium, including electron collisions, is carried out. Section F discusses further improvements of the theory which are presently under way or will be required in the future. Finally, Section G deals with the basic spectroscopic observational tests that must be undertaken to investigate whether the theory can be regarded as quantitatively reliable for future astronomical use.

B. Theory of Radiation-Driven Winds

Hot stars have an extremely intense radiation field. The absorption of this radiation by UV metal lines in the outer atmospheric layers yields an absorption of outward-directed momentum and, consequently, an outward accelerating force that exceeds that of gravity by a large factor. It is well known that this radiative line force is sufficient to initialize and maintain stellar winds (Lucy and Solomon, 1970; Abbott, 1979). This connection is displayed in Figure 4-6, which shows that almost all massive OB stars with a luminosity larger than that corresponding to an evolutionary track with initial mass of $M = 15\ M_\odot$ suffer from mass loss. This strikingly coincides with the domain of self-initiating radiation-driven winds (Abbott, 1979) that is obtained from a comparison of radiative-line forces with photospheric gravities. Obviously, for stars with initial masses $M > 15\ M_\odot$, stellar winds are already present at the main sequence and are maintained throughout the stellar lifetime.

In the following, we will outline the basic ideas of the radiation-driven wind theory and discuss its predictions for the observable quantities such as mass-loss rate $\dot{M}$ and terminal velocity v_∞. We will proceed in three steps. First, we discuss a very simplified version.

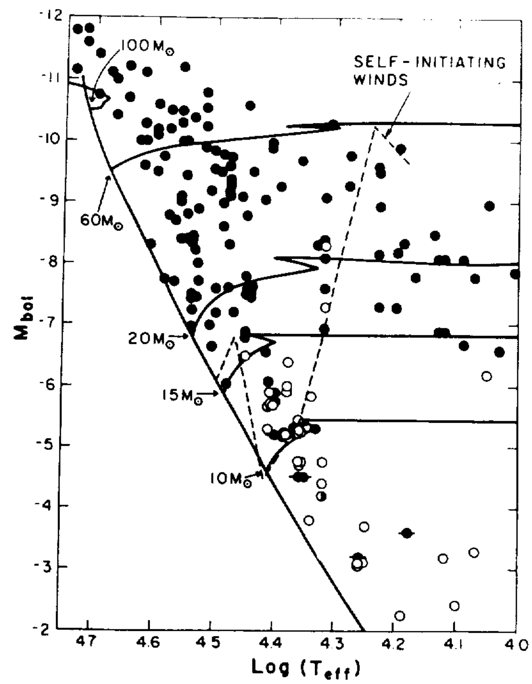


Figure 4-6. The HR diagram of massive stars (from Abbott, 1982b): o = objects with mass loss, o = mass loss, but rapid Be-type rotators, and o = no mass loss. The domain of self-initiating radiation-driven winds (from Abbott, 1979) is given by the theoretical ZAMS and the dashed line.

However, this already yields the basic relations. Second, we outline the more realistic version by Castor et al. (1975a, referred to in equations as CAK) and compare its best computations (Abbott, 1982) with observations. Third, we introduce the modified Castor et al. theory by Pauldrach et al. (1986) and discuss its consequences on the wind dynamics.

1. Supersimplified Theory. The physical situation is sketched in Figure 4-7. A spherical, differentially expanding shell of geometrical thickness Δr and optical thickness $\tau_s(r)$ and of density ρ (r) is irradiated by the underlying stellar photosphere. The simplifying assumptions in this step are:

a. Only radially streaming photons interact with the wind.

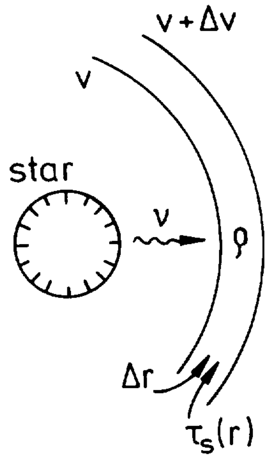


Figure 4-7. A differentially expanding radiation-driven spherical shell in the wind (see text).

b. Only strong lines that absorb in their frequency bandwidth Δv_i all the radiation ($\tau_s >> 1$) drive the wind.

c. The velocity gradient due to the differential expansion is so large that the line bandwidth across the shell is determined by the Doppler formula (Sobolev approximation):

$$\Delta \nu_i = \frac{\Delta v}{c} \nu_i \; .$$

The latter assumption means that $\Delta v_i >> \Delta v_D$ (the thermal Doppler width) and is equivalent to $\Delta v >> v_{th}$ (thermal velocity of the ion producing the line) or $v_{th}/(dv/dr) \ll l_{wind}$ (the characteristic scale length in the wind). Assumptions a and b will be discussed in Sections 2 and 3, respectively.

On the basis of these assumptions, radiative line acceleration g_{rad} can be easily computed:

$$g_{rad} = \frac{\text{absorbed momentum}}{\Delta t \; \Delta m}$$

$$\Delta m = 4\pi r^2 \rho \, \Delta r$$

$$\frac{\text{abs. mom.}}{\Delta t} = \underbrace{\frac{L}{C}}_{\substack{\text{Total mom.}\\ \text{provided by}\\ \text{star}}} \quad \underbrace{\frac{L_\nu \, \Delta \nu_i}{L}}_{\substack{\text{Fraction absorbed}\\ \text{by strong line } i \text{ of}\\ \text{frequency } \nu_i\\ \text{and width } \Delta\nu_i}}$$

Frequency bandwidth $\Delta \nu_i$ of the line can be replaced by the Doppler formula (assumption c), which, after summing over all lines, then yields:

$$g_{rad} = \frac{L}{c^2} \underbrace{\sum_i \frac{L_\nu \nu_i}{L}}_{: \, = N_{eff}} \frac{1}{4\pi r^2 \rho} \frac{dv}{dr} \tag{4-5}$$

We see that g_{rad} is proportional to the luminosity and to the velocity gradient. The dimensionless quantity N_{eff} results from the sum over all lines and is interpreted as the number of effectively acting strong lines. The use of Equation (4-5) allows the approximative formulation of the (stationary) equation of motion of the stellar wind fluid in the supersonic region:

$$\underbrace{v \frac{dv}{dr}}_{\text{inertia}} = \underbrace{\frac{L}{c^2} N_{eff} \frac{1}{4\pi r^2} \frac{dv}{dr}}_{g_{rad}} \underbrace{- \frac{GM\,(1-\Gamma)}{r^2}}_{g \, - \, g_{\text{Thomson}}} \; . \tag{4-6}$$

Note that, in principle, we would have to include the terms arising from the gradient of the gas pressure. This means that, in the case of an isothermal cool wind ($T_e \approx T_{eff}$) with sound speed v_s, we would have to add

$$- \frac{v_s^2}{v^2} v \frac{dv}{dr}$$

on the left and

$$+ \frac{2v_s^2}{r}$$

on the right side of Equation (4-6).

However, because we want to restrict our approximative treatment to the supersonic region ($v >> v_s$), which, according to the observations, comprises almost the entire wind, we can easily neglect the additional term on the left side. Moreover, for massive OB stars with a cool wind, the inequality

$$\frac{2v_s^2}{r} << \frac{GM\,(1-\Gamma)}{r^2}$$

holds up to more than 100 stellar radii in most of the cases. This means that the additional term on the right side can also be neglected until very far out in the wind. Thus, to a first approximation, the wind can be treated as a collection of noninteracting particles, as in Equation (4-6).

If we combine Equation (4-6) with the equation of continuity,

$$\dot{M} = 4\pi\, r^2 \rho\, \mathrm{v} = \text{const.} \quad , \qquad (4\text{-}7)$$

we obtain

$$\dot{M} = \frac{L}{c^2}\, N_{\mathrm{eff}}\,(1-\epsilon) \quad , \qquad (4\text{-}8)$$

where the ϵ is given by

$$\epsilon = \frac{\mathrm{GM}\,(1-\Gamma)\,4\pi\rho}{\mathrm{dv/dr}}\;\frac{c^2}{L\,N_{\mathrm{eff}}} \quad . \qquad (4\text{-}9)$$

From Equation (4-8), we see that $\epsilon(\mathrm{r})$ must be constant if N_{eff} is constant. The value of ϵ obviously indicates how much of the radiative momentum arising from the stellar luminosity is transferred to the stellar wind flow. In this supersimplified step, we do not intend to determine ϵ by means of the theory. Instead, we intend to derive it from the observed terminal velocities.

Inserting Equations (4-7) and (4-8) into Equation (4-5) gives a simplified expression for g_{rad}:

$$g_{\mathrm{rad}} = \frac{1}{1-\epsilon}\;\mathrm{v}\;\frac{\mathrm{dv}}{\mathrm{dr}} \quad . \qquad (4\text{-}10)$$

This means that the functional dependence of the line acceleration on the velocity field is the same as for the inertia of the material that must be accelerated. As a consequence, the velocity is determined completely by the gravity:

$$\frac{\epsilon}{1-\epsilon}\;\mathrm{v}\;\frac{\mathrm{dv}}{\mathrm{dr}} = \frac{GM\,(1-\Gamma)}{r^2} \quad . \qquad (4\text{-}11)$$

Equation (4-11) can be easily integrated yielding:

$$\mathrm{v}^2\,(r/R_*) = \frac{1-\epsilon}{\epsilon}\;\mathrm{v}_{\mathrm{esc}}^2\;(1-R_*/r) \quad ,$$
$$\mathrm{v}_{\mathrm{esc}} = \frac{2\,GM\cdot(1-\Gamma)}{R_*} \quad . \qquad (4\text{-}12)$$

$\mathrm{v}_{\mathrm{esc}}$ is the escape velocity from the stellar surface of radius R_*. The consequence of Equation (4-12) is that terminal velocity v_∞ is predicted to be proportional to $\mathrm{v}_{\mathrm{esc}}$:

$$\mathrm{v}_\infty = \left(\frac{1-\epsilon}{\epsilon}\right)^{1/2}\;\mathrm{v}_{\mathrm{esc}} \quad . \qquad (4\text{-}13)$$

This theoretical prediction is clearly confirmed by the observations as displayed in Figure 4-8, which indicates that the idea of radiative acceleration is probably useful. To the first order, the observations suggest a relation $\mathrm{v}_\infty = 3\,\mathrm{v}_{\mathrm{esc}}$ (Abbott, 1978), although a large scatter and a weak additional systematic behavior is present (see Abbott, 1982a). The observed proportionality constant of 3 by Equation (4-13) indicates that ϵ is about 0.1 and to the first order independent of the stellar parameters (at least for luminous OB stars with galactic metallicity). If this is true, then by Equation (4-8), the theory predicts that the observed mass-loss rates should be proportional to the stellar luminosities.

Figure 4-9 shows that in fact a strong correlation of $\dot{M}$ and L is observed for massive OB stars, indicating that the basic concept of radiation-driven winds is correct. However, the observed luminosity dependence is significantly greater and yeilds $\dot{M} \sim L^{1.6}$ (Garmany and Conti, 1984). In the next step, we will show that this can be explained by the refined theory.

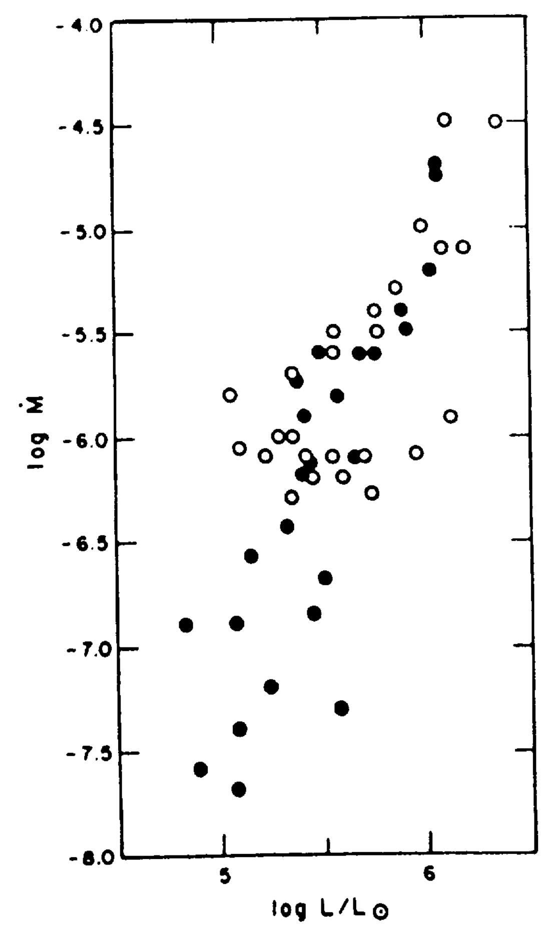


Figure 4-8. The observed correlation between mass-loss rate and luminosity (from Garmany and Conti, 1984).

2. The Castor et al. Theory. The simple theory developed in the foregoing section has a severe disadvantage. It allows for only strong lines and assumes $N_{eff}(r)$ = constant (i.e., the number of effectively acting strong lines to be constant in the wind). However, this cannot be completely correct because, for an arbitrary line i, the optical thickness of the spherical differentially expanding shell $\tau_s^i(r)$ must decrease outward with decreasing density (if no significant changes in ionization and excitation occur). Since the fraction of luminosity absorbed by a line of optical thickness τ_s and bandwidth Δv_i is given by:

$$\frac{L_\nu \, \Delta \nu}{L} \left(1 - e^{-\tau_s^i} \right) ,$$

the contribution g^i_{rad} of this line to the radiative acceleration is described by

$$g^i_{rad} = \frac{L}{c^2} \frac{L_\nu \nu_i}{L} \left(1 - e^{-\tau_s^i} \right) \frac{1}{4\pi r^2 \rho} \frac{dv}{dr} . \qquad (4\text{-}14)$$

For a strong line ($\tau_s >> 1$) Equation (4-14) reduces to Equation (4-5). However, for weak lines ($\tau_s << 1$), the expansion of the exponential yields:

$$g^i_{rad} = \frac{L}{c^2} \frac{L_\nu \nu_i}{L} \frac{\tau_s^i}{4\pi r^2 \rho} \frac{dv}{dr} . \qquad (4\text{-}15)$$

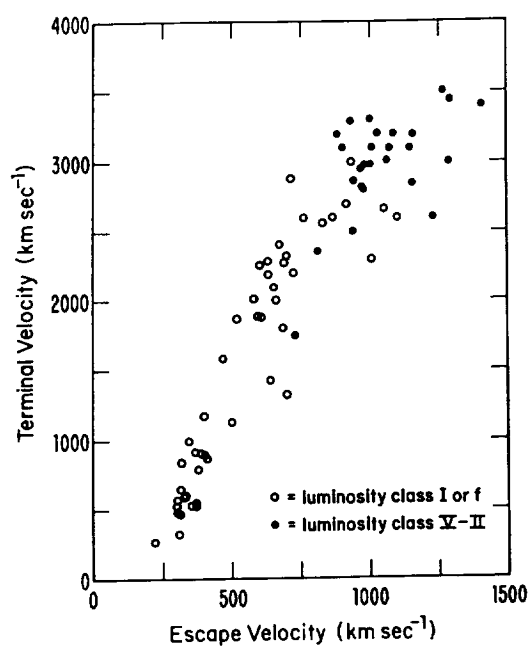


Figure 4-9. The observed correlation between terminal velocity and photospheric escape velocity (from Abbott, 1982a).

The theory of Castor et al. (1975a) now makes the following assumptions:

a. Only radially streaming photons interact with the wind (as in Section B.1).

b. Strong ($\tau_S >> 1$ *and* weak ($\tau_S << 1$) lines drive the wind.

c. Same as in Section B.1.

To decide whether a line is strong or weak and to compute the radiative force for weak lines, it is necessary to estimate $\tau_s^i(r)$. In rapidly expanding atmospheres, this is rather easy. The line absorption coefficient χ_ν^i is given by:

$$\chi_\nu^i = \varkappa(r)\,\varphi\,(x-v(r)/v_{th}) \ ,$$

$$\varkappa(r) = \frac{\pi e^2}{m\,c}\,\frac{1}{\Delta\nu_D}\,g_l f_{lu} \left(\frac{n_l(r)}{g_l} - \frac{n_u(r)}{g_u}\right) , \tag{4-16}$$

$$x = \frac{\nu-\nu_i}{\Delta\nu_D} \ ,$$

$$\varphi(x) = (\pi)^{-1/2}\,e^{-x^2} \ .$$

If we now choose the frequency ν in the line such that $\nu = \nu_i + \Delta\nu_D\, v(r)/v_{th}$, then the radial optical depth integral,

$$\tau_s^i = \int_{-\Delta r/2}^{\Delta r/2} \chi_\nu^i\,(\tilde{r})\,d\tilde{r} \ ,$$

is dominated by the strong peak of the profile function $\varphi\,(x-v(r)/v_{th})$ at $\tilde{r} = r$. If we now assume $\varkappa[r)$ to vary slowly over Δr (see assumption c), then we have:

$$\tau_s^i \approx \varkappa(r) \int_{-\Delta r/2}^{\Delta r/2} \varphi\,(x-v\,(\tilde{r})/v_{th})\,d\tilde{r} \ .$$

Taking into account the normalization of φ and the large velocity gradient, this yields approximately

$$\tau_s^i(r) \approx \varkappa(r)\,\frac{v_{th}}{dv/dr} \ .$$

Introducing the line strength k_i by:

$$\varkappa(r) = k_i\,\rho(r) \tag{4-17}$$

with k_i to the first order density and depth independent (this is of course true only as long as $n_l/ \ \approx$ constant), we finally obtain:

$$\tau_s^i(r) = k_i\,v_{th}\,\frac{\rho(r)}{dv/dr} \ . \tag{4-18}$$

This means that the optical thickness increases with the reciprocal velocity gradient. Inserting Equation (4-18) into Equation (4-15), one obtains for a weak line ($\tau_s << 1$):

$$g_{rad}^i = \frac{L}{c^2}\,\frac{L_\nu\,\nu_i}{L}\,k_i\,v_{th}\,\frac{1}{4\pi\,r^2}\,\frac{1}{r^2} \ ,$$

with the same depth dependence as for Thomson scattering. Consequently, the total radiative acceleration in the Castor et al. (1975a) theory contains two terms:

$$g_{\rm rad}^{\rm CAK} = \frac{L}{c^2} \frac{1}{4\pi r^2} \left(N_s(r) \frac{1}{\rho} \frac{dv}{dr} + N_W(r) \langle k_i \rangle v_{th} \right) , \tag{4-19}$$

$$N_s(r) = \sum_{\substack{\text{strong} \\ \text{lines}}} \frac{L_\nu \nu_i}{L} ,$$

$$N_W(r) \langle k_i \rangle = \sum_{\substack{\text{weak} \\ \text{lines}}} \frac{L_\nu \nu_i}{L} k_i .$$

The stratification of $N_s(r)$ and $N_W(r) \langle k_i \rangle$ is now computed from the statistical distribution of line strengths k_i. Taking a large sample of lines that contribute to the line force (see below) and calculating density-independent line strengths k_i under wind conditions, one finds that the line-strength distribution function $dN(k)$ can be described by:

$$d N(k) = N_0 (1-\alpha) (k/s_E)^{\alpha-2} \, dk/s_E , \tag{4-20a}$$

or

$$N(k \geqslant k_0) = N_0 (k_0/s_E)^{\alpha-1} . \tag{4-20b}$$

Here $N(k \geqslant k_0)$ denotes the number of lines with a strength k larger than a certain limit k_0. (s_E is the Thomson scattering coefficient divided by ρ.) N_0 is the number of all lines, strong and weak, that really contribute to the line force. The parameter α then gives the steepness of this function. Note that α is not a free parameter, but is determined from real line calculations in the wind. An example in which the 250000 lines of Abbott's (1982a) list were used and calculated by J. Puls is given in Figures 4-10a and 4-10b.

Then, from

$$\tau_s^i(r) \quad \begin{matrix} > 1 \text{ strong line} , \\ < 1 \text{ weak line} \end{matrix}$$

the limit $k_0(r)$ defines at every depth the division between strong and weak lines and can be easily computed to be:

$$k_0(r) = \frac{dv/dr}{\rho \, v_{th}} . \tag{4-21}$$

Using Equations (4-21) and (4-20b), we obtain immediately for the number of strong lines

$$N_s(r) = N_o \left(\frac{1}{s_E \rho \, v_{th}} \frac{dv}{dr} \right)^{\alpha-1} . \tag{4-22}$$

The average over the weak lines in the second term of Equation (4-19) is given by ($0 < \alpha < 1$ is assumed):

$$N_W(r) \langle k \rangle = \int_{s_E}^{k_0} k \, dN(k)$$

$$= N_0 \frac{1-\alpha}{\alpha} \left[\left(\frac{k_0}{s_E} \right)^{\alpha-1} k_0 - s_E \right] .$$

The second term in the parentheses can be neglected and, with Equation (4-21), yields:

$$N_W(r) \langle k \rangle \approx N_0 \frac{1-\alpha}{\alpha} \left(\frac{1}{s_E \rho \, v_{th}} \frac{dv}{dr} \right)^{\alpha-1} \frac{1}{\rho v_{th}} \frac{dv}{dr} . \tag{4-23}$$

This, together with Equations (4-19) and (4-22), gives the final expression for the line force in the approximation,

$$g_{\rm rad}^{\rm CAK} = \frac{L}{c^2} \frac{N_0}{4\pi r^2} s_E \, v_{th} \left(\frac{dv/dr}{s_E \rho \, v_{th}} \right)^{\alpha} \frac{1}{\alpha} : \tag{4-24}$$

$0 < \alpha < 1$. Note that, in the original Castor et al. paper, the line force was not derived in this

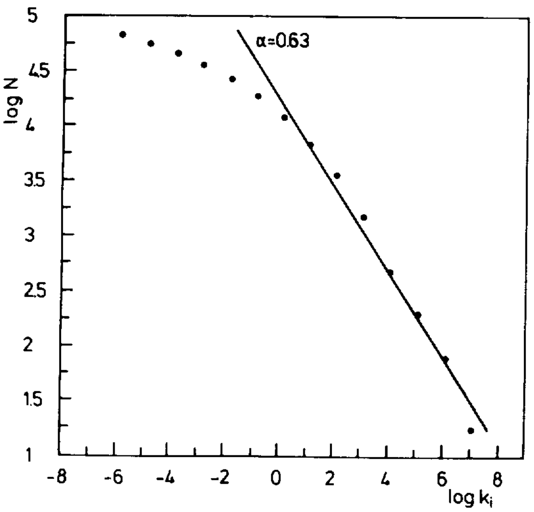


Figure 4-10a. Logarithm of line-strength distribution function versus logarithm of line strength k_i for a typical hot-star wind model. Note the power-law dependence for log $k \geq 0$, which represents those lines really contributing to the force.

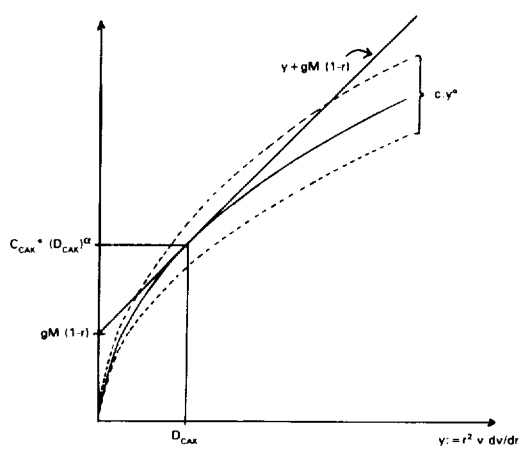


Figure 4-10b. The possible solutions for the velocity gradient r^2 v dv/dr as obtained from Equation (4-27a) (see text).

way, but was obtained by an empirical fit of the force as a function of atmospheric depth. For discussion, see also Abbott (1982a).)

Now we have a nonlinear dependence on the velocity gradient, introduced by the line-distribution steepness parameter α. (Note that $\alpha = 1$ corresponds to the strong line limit.) Although the theory in detail is greatly complicated by the nonlinearity of Equation (4-18), relatively simple estimates for $\dot{M}$ and v_∞ are possible. The simplified equation of motion, instead of Equation (4-7), now reads:

$$v \frac{dv}{dr} = \frac{L}{c^2} N_o \frac{1}{4\pi r^2} s_E v_{th} \left(\frac{1}{s_E \rho v_{th}} \frac{dv}{dr} \right)^\alpha - \frac{GM(1-\Gamma)}{r^2} \quad . \tag{4-25}$$

Eliminating ρ by using Equation (4-6), the continuity equation, and defining

$$C = \frac{L}{c^2} \frac{N_o}{4\pi} s_E v_{th} \left(\frac{4\pi}{s_E v_{th} \dot{M}} \right)^\alpha , \tag{4-26}$$

we obtain

$$r^2 v \frac{dv}{dr} = C \left(r^2 v \frac{dv}{dr} \right)^\alpha - GM(1-\Gamma) \quad . \tag{4-27a}$$

The reader may note that the complete equation of motion, including a variable temperature structure, reads:

$$r^2 v \left(1 - \frac{v_s^2}{v^2} \right) \frac{dv}{dr} = C \left(r^2 v \frac{dv}{dr} \right)^\alpha - GM(1-\Gamma) + 2 v_s^2 r - r^2 \frac{dv_s^2}{dr} \tag{4-27b}$$

and was solved correctly by Castor et al. (1975a) and Pauldrach et al. (1986). For constant temperature, however, our simplified

Equation (4-27a) is justified for the entire supersonic flow by the discussion following Equation (4-6).

Then, Equation (4-27) can be fulfilled at every depth only if

$$r^2 \, v \, \frac{dv}{dr} = D = \text{const.} \qquad (4\text{-}28)$$

The situation is sketched in Figure 4-10b, which shows that either two, one, or no solutions are possible for D if the constant C (which contains $\dot{M}$) and the wind flow are unrelated. However, if the wind is purely radiatively driven, then $\dot{M}$ and the wind flow are related and consequently only one stationary solution should be possible. In this case, both constants C and D are selfconsistently determined by Equations (4-27) and (4-28) and the additional condition (see Figure 4-10b):

$$C \, \alpha \, D^{\alpha-1} = 1 \quad . \qquad (4\text{-}29)$$

This yields:

$$D = \frac{\alpha}{1-\alpha} \, GM \, (1-\Gamma) \quad ,$$

$$(4\text{-}30)$$

$$C = \frac{1}{\alpha} \left(\frac{\alpha}{1-\alpha} \right)^{1-\alpha} (GM \, (1-\Gamma))^{1-\alpha} \quad .$$

Integration of Equation (4-28) with the value for D given by Equation (4-30) leads to:

$$v^2 \left(\frac{r}{R_*} \right) = \frac{\alpha}{1-\alpha}$$

$$v^2_{esc} \left(1 - \frac{R_*}{r} \right) \quad , \qquad (4\text{-}31)$$

$$v_\infty = \left(\frac{\alpha}{1-\alpha} \right)^{1/2} v_{esc} \quad .$$

($v(R_*) << v(r)$ is assumed implicitly as for Equation (4-12).)

From Equations (4-27) and (4-30), we obtain:

$$\dot{M} \; L^{1/\alpha}{}_* \, M^{(1-1/\alpha)} \quad . \qquad (4\text{-}32)$$

We note that the approximative solutions obtained here are identical to those given originally by Castor et al. (1975a), but the way in which they have been derived is different. (Castor et al. used the singularity and regularity conditions at the critical point that were in any case required for the correct numerical theory and then arrived at these simplified expressions.)

We see that the luminosity dependence of $\dot{M}$ has now become stronger as observed (see Figure 4-8). (Note that $0 < \alpha < 1$, with realistic values being between 0.5 and 0.7.) On the other hand, the linear relation between v_∞ and v_{esc} still remains, as is also observed (see Figure 4-9). However, as far as the proportionality constants for both relations are concerned, the theory has severe problems. This has been pointed out by Abbott (1982a), who has thus far published the most realistic calculations within the framework of the Castor et al. (1975a) theory. His calculations were made using a list of 250000 lines that is almost complete for the elements from H to Zn in the ionization stages I to VI. The main results are given in Figures 4-11 and 4-12. They show that the slope of the log $\dot{M}$ /log L relation predicted by the theory now agrees with the observations. However, the calculated mass-loss rates are too large by a factor of 3. Although this is probably not too disturbing, in view of the observational errors in the $\dot{M}$ determination, the theory completely fails to predict the observed terminal velocities. It predicts $v_\infty/v_{esc} \approx 1.2$, whereas a ratio of 2 to 3 is observed. This discrepancy is extremely alarming since terminal velocities can usually be measured with high precision (+10 percent). We therefore must conclude that, despite the enormous effort (250000 lines), the theory at this stage is still not satisfactory. However, as will be shown in the next section, a further significant improvement is possible.

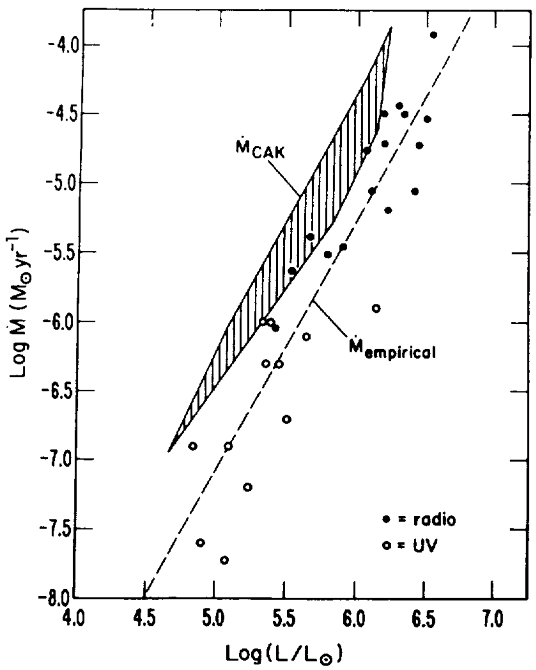


Figure 4-11. Observed mass-loss rate versus luminosity compared with the prediction of the Castor et al. (1975a) theory (from Abbott, 1982).

3. Improved Theory of Radiation-Driven Winds. The assumption of only radially streaming photons made in the foregoing steps will be correct only at very large distances from the star, where it appears approximately as a point source. Castor et al. (1975a) realized this, but neglected it for reasons of simplicity. On the other hand, Pauldrach et al. (1986) investigated this effect and found it to be extremely important. Their improved theory is characterized by the following major assumptions:

a. Nonradial photons are also allowed to contribute to the line force.

b. Same as in Section B.2.

c. Same as in Section B.1.

As a consequence of a, the photospheric finite cone angle, which irradiates the differentially expanding shell, must be considered (see Figure 4-13) for the interaction with the wind. This means that nonradial rays also intersect the shell. To calculate their optical thickness, we must replace dv/dr by $d(\mu v)/ds = (\mu^2\, dv/dr + (1-\mu^2)\, v/r)$ in Equation (4-18), where μ is the cosine of the angle between the radial and a nonradial ray. Consequently, the optical thickness becomes angle-dependent:

$$\tau_s^i\,(r,\mu) = k_i\, v_{th}\, \rho\, (\mu^2\ dv/dr + (1-\mu^2)\ v/r)^{-1}\ .$$

(In addition, for the calculation of $\Delta\nu_i$, dv/dr has to be replaced by $d(\mu v)/ds$.) As a result, the contribution of a single line to the line acceleration g^i_{rad} now reads:

$$g^i_{rad} = \frac{L}{c^2}\,\frac{L_\nu\,\nu_i}{L}\,\frac{1}{4\pi\, r^2 \rho}\ * \ ,$$

$$\frac{2}{1-\mu_*^2}\int_{\mu^*}^{1}\left(\left[\mu^2\,\frac{dv}{dr} + (1-\mu^2)\,\frac{v}{r}\right]\right. \tag{4-33}$$

$$\left.\left[\,1-e^{-\tau_s(r,\mu)}\right]\right)\mu d\mu\ .$$

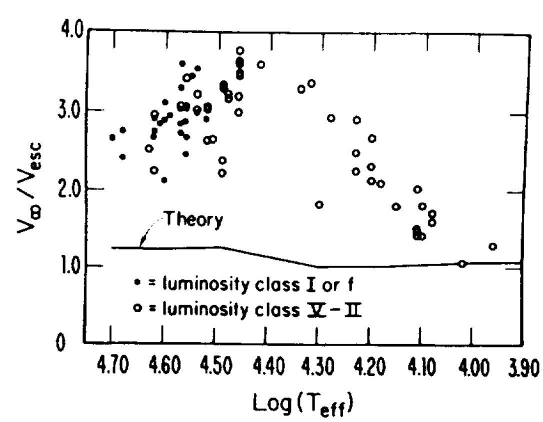


Figure 4-12. The observed ratio of v_∞/v_{esc} as a function of log T_{eff}. The prediction of the Castor et al. (1975a) theory is also indicated (from Abbott, 1982a).

Equation (4-33) can be used to derive a correction factor (CF) for the Castor et al. (1975a) radiative force and gives:

$$g_{\rm rad} = g_{\rm rad}^{\rm CAK}\, CF\,(r)\ ,$$

$$CF\,(r) = \frac{2}{1-\mu_*} \int_{\mu_*}^{1} \left(\frac{\mu^2\, {\rm dv/dr} + (1-\mu^2)\, {\rm v/r}}{\rm dv/dr} \right)^{\alpha} \mu d\mu \ . \qquad (4\text{-}34)$$

The introduction of this CF into the full hydrodynamic equations (which also contain the pressure terms discussed on page 5 when treated numerically) complicates the numerical and mathematical treatment of the theory significantly. However, as shown by Pauldrach et al. (1986) efficient algorithms can be found that make the solution of the problem possible. (Similar calculations were recently made available to us by Friend and Abbott (1986).) Figure 4-14 shows the stratification of the CF for the case of a typical O5 star resulting from the improved calculations. As to be expected from simple geometrical reasons, CF becomes smaller than unity in the deeper layers in the wind (to be exact, for the region where dv/dr > v/r). Because the mass-loss rate is fixed in the deeper layers, we expect a smaller $\dot{M}$ to result from the introduction of CF. Very interestingly, in the outer layers (dv/dr < v/r) CF

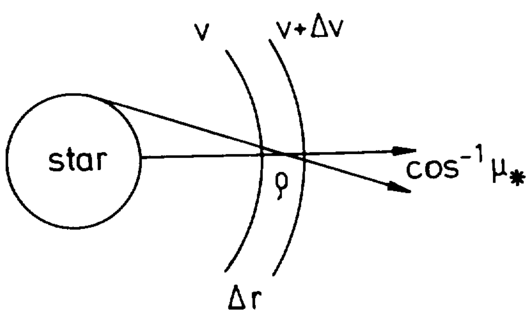


Figure 4-13. Photospheric finite cone angle that irradiates the expanding envelope.

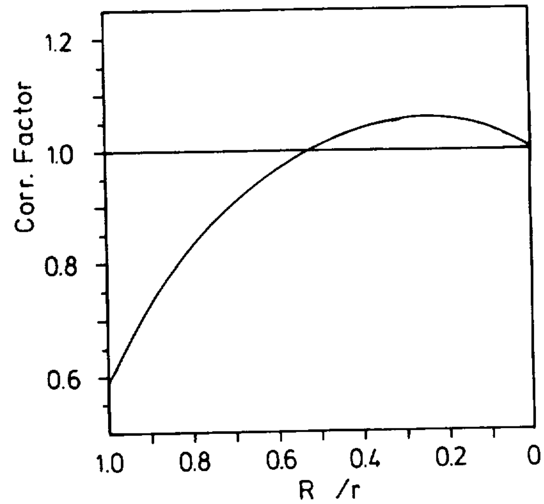


Figure 4-14. Finite cone-angle correction factor for a typical O5 star as a function of the reciprocal radius (from Pauldrach et al., 1986).

becomes larger than unity and then approaches it from above. This means that here the acceleration is increased. Since the density of the material to be accelerated is now smaller because of the lower mass-loss rate in the deeper layers, much higher velocities are obtained. This means that the wind dynamics are changed significantly, and much better agreement with the observations is achieved. This will be discussed in the following section.

C. Results of Improved Theory

1. Galactic OB Stars. Pauldrach (1986) has applied the improved theory described in the foregoing section to a representative sample of galactic OB stars. For the calculations, they have chosen three O V stars, two evolved O stars, and two B supergiants, all with well-known wind parameters M and v_∞ and not too uncertain stellar parameters ($T_{\rm eff}$, log g, and $L/L_\odot$).

Table 4-3 lists the objects, together with their stellar parameters and the observed and calculated values of $\dot{M}$ and v_∞. The stellar parameters have been somewhat adjusted within the allowed limits of errors to demonstrate that, in principle, the observations can be fully matched by the theory. An improved determination of stellar parameters, which is at hand with present-day observational techniques (see Kudritzki and Hummer, 1986; Kudritzki, 1985) will, of course, lead to more precise constraints on the theory. (See also Section G.)

Strikingly, the agreement between theory and observations is very good for all the objects in Table 4-3. The observed mass-loss rates and terminal velocities are reproduced by the theory not only for the O main-sequence stars and the evolved O stars, but also for the extreme prototype of stars showing mass loss—P Cygni, which has a very high $\dot{M}$ and an extremely low v_∞. This reasonable agreement between the improved selfconsistent theory and the observations in both $\dot{M}$ and v_∞ over a rather large domain in the HR diagram indicates that the concept of radiation-driven winds is certainly very promising for a quantitative description of the time-averaged stationary wind properties of hot luminous stars.

2. O Stars in the Magellanic Clouds. With respect to the evolution of galaxies, it is extremely important to study the influence of metallicity on the strength of stellar winds. For this purpose, the Magellanic Clouds are an ideal laboratory, since the analyses of H II region emission-line spectra indicate a significant underabundance of metals in the Large Magellanic Cloud (LMC), as well as in the Small Magellanic Cloud (SMC) (see for instance, Dufour, 1984). Consequently, Kudritzki et al (1986) have applied their improved theory of radiation-driven winds to O stars in the LMC and SMC. For their calculations, they adopted the following metal abundances:

$$Z_{\rm LMC} = 0.28 \cdot Z_{\rm Gal},\ Z_{\rm SMC} = 0.1 \cdot Z_{\rm Gal} \ .$$

With these metal abundances, they calculated new line forces, again using the sample of 250000 lines provided by Abbott (1982a) and a variety of wind models. As an example, results for a typical O5 V star ($T_{\rm eff}$ = 45000 K, log g = 4.0, $R/R_\odot$ = 12) are given in Table 4-4.

The improved theory predicts a significant decrease with metal abundance for both $\dot{M}$ and v_∞ . The effect amounts to 900 km/s in v_∞ and a factor of 3 in $\dot{M}$.

It is extremely interesting to compare such calculations with observations. Fortunately, several studies of O-star winds have been carried out with the IUE satellite (Bruhweiler et

Table 4-3
Mass-Loss Parameters in OB-Type Stars

Star	Spectral Type	$T_{\rm eff}$ (10^3 K)	log g (cgs)	log $L/L_\odot$	$10^{-6}\ M_\odot$/yr $\dot{M}_{obs}$	$\dot{M}_{calc}$	km/s v_∞^{obs}	v_∞^{calc}
P Cyg	B Ia	18.0	2.0	5.64	20–30	29	400	395
ϵ Ori	B0 Ia	28.5	3.25	5.91	3.1	3.3	2010	1950
ζ Ori	O9.5 I	30.0	3.45	5.79	2.3	1.9	2290	2274
9 Sgr	O4(f) V	50.0	4.10	5.95	4.0	4.0	3440	3480
HD 48099	O6.5 V	39.0	4.00	5.40	0.63	0.64	3500	3540
HD 42088	O6.5 V	40.0	4.05	4.89	0.13	0.20	2600	2600
λ Cep	O6ef	42.0	3.7	5.90	4.0	5.1	2500	2500

Table 4-4
Calculated Wind Parameters for a Typical O5 V Star in the Galaxy, LMC, and SMC

	v_∞ km/s	$\dot{M}$ $10^{-6}\ M_\odot$/yr	$\frac{\dot{M} v_\infty C}{L}$
Z_{Gal}	3350	2.12	0.66
Z_{LMC}	2900	1.35	0.36
Z_{SMC}	2435	0.72	0.16

al., 1982; Hutchings, 1982; Garmany and Conti, 1985), mainly in the low-resolution mode. Figure 4-15 shows a significant result found by Garmany and Conti (1985) for O stars of spectral type O3 to O9 and luminosity class V to III in the Galaxy, LMC, and SMC. The terminal velocities v_∞ form an inclined band as a function of T_{eff} for stars in all three galaxies. However, on the average, v_∞ in the LMC and the SMC are about 600 and 1000 km/s lower than in the Galaxy. Therefore, a general trend

$$v_\infty^{SMC} < v_\infty^{LMC} < v_\infty^{Gal}$$

can be inferred from the observations.

Kudritzki et al. (1986) have investigated this trend in more detail. They calculated wind models along evolutionary tracks, computed recently by Pylyser et al. (1985) for SMC, LMC, and Galaxy. Both wind models and tracks used the same abundances as given above. The results are shown in Figure 4-16 and reveal the following. Obviously the evolutionary tracks for O stars lead to an inclined band in the (v_{esc}, T_{eff})-plane that is defined by the ZAMS and 40 $M_\odot$ track. Since v_∞ and v_{esc} are related, we expect similar bands in the (v_∞, log T_{eff})-plane, which is indeed true as shown in the lower part of Figure 4-16. However, due to the different metallicity in each galaxy (see Table 4-4), the theoretical bands are shifted relatively in height. The relative positions of the (v_∞, log T_{eff})-bands are shown in Figure 4-17. This explains nicely the observations given in Figure 4-15. The location of the individual objects of each galaxy in the lower part of Figure 4-16 also demonstrates that the observed v_∞ values are well represented by the improved theory.

The mass-loss rates along the evolutionary tracks in each galaxy are displayed in Figure 4-18a. Although, individually, for tracks of the same initial mass, the mass-loss rates are clearly smaller in the Magellanic Clouds, it will be hard to disentangle this effect because, at higher luminosities, the curves cross each other. This explains why no clear observational effects have been found by Garmany and Conti (1985) with respect to mass-loss rates. However, by means of detailed non-LTE quantitative spectroscopy that yields the stellar parameters (T_{eff}, log g, L, M, *and* abundances) with high precision for individual objects, this will be possible. Such work is presently under way, and first results will be forthcoming soon (Gehren et al., 1986).

The absolute values of the computed mass-loss rates (at least for the Galaxy) are in good agreement with the observations. This is

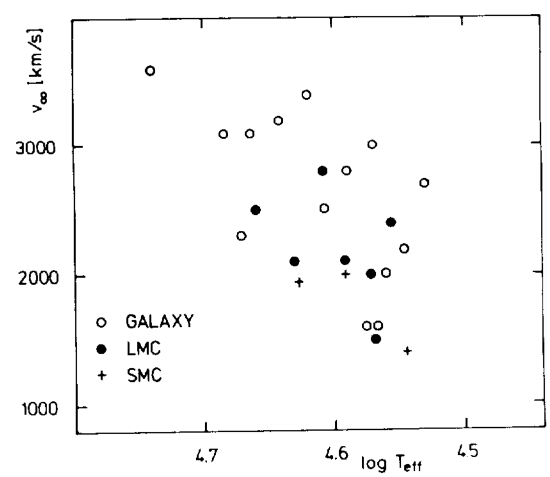


Figure 4-15. Observed terminal velocities as a function of effective temperature for O stars in the Galaxy, LMC, and SMC. The data are taken from Garmany and Conti (1985).

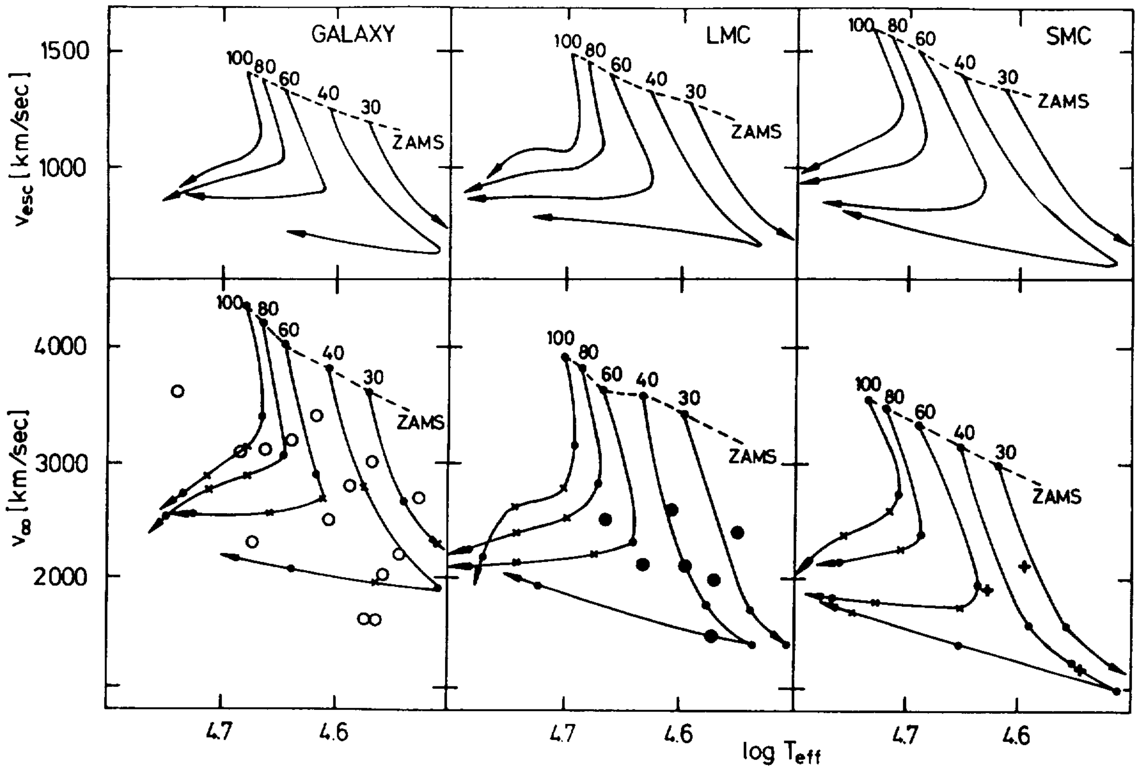


Figure 4-16. Top: The change of surface escape velocity during the evolution of massive stars for the Galaxy, LMC, and SMC. The curves are computed from the evolutionary tracks of Pylyser et al. (1985) and are labeled by their initial mass at the ZAMS. Bottom: Terminal velocity computed with the improved stellar-wind theory along the same tracks as in the corresponding upper diagram. Observed values have been added for comparison from Figure 4-15. For a detailed discussion, see text (from Kudritzki, et al., 1986).

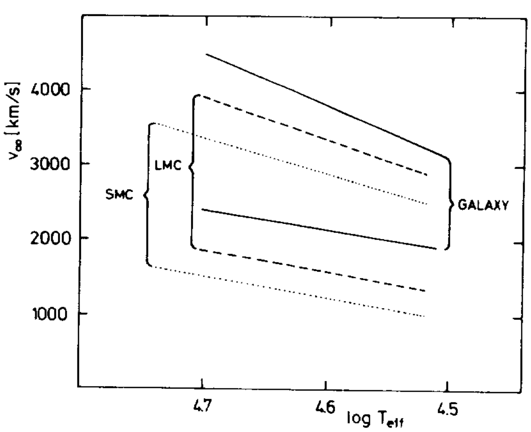


Figure 4-17. The v_∞-bands as a function of T_{eff} for the Galaxy, LMC, and SMC as predicted by the improved theory (from Kudritzki et al., 1986).

demonstrated by Figure 4-18b, which also indicates that part of the large scatter in $\dot{M}$ might come from stars with different M evolving along different tracks.

In general, the improved theory predicts winds that are significantly weaker in velocity and density if the metal abundance is reduced to values suggested for the LMC and the SMC. From the fact that the observed decrease of the terminal velocity is well reproduced by the calculations, we conclude that the concept of radiation-driven winds appears to function correctly at low metallicities as well. On the other hand, the results can also be interpreted as an indirect proof of the overall low metallicity in the LMC and SMC if one is willing to accept the wind theory as a priori correct.

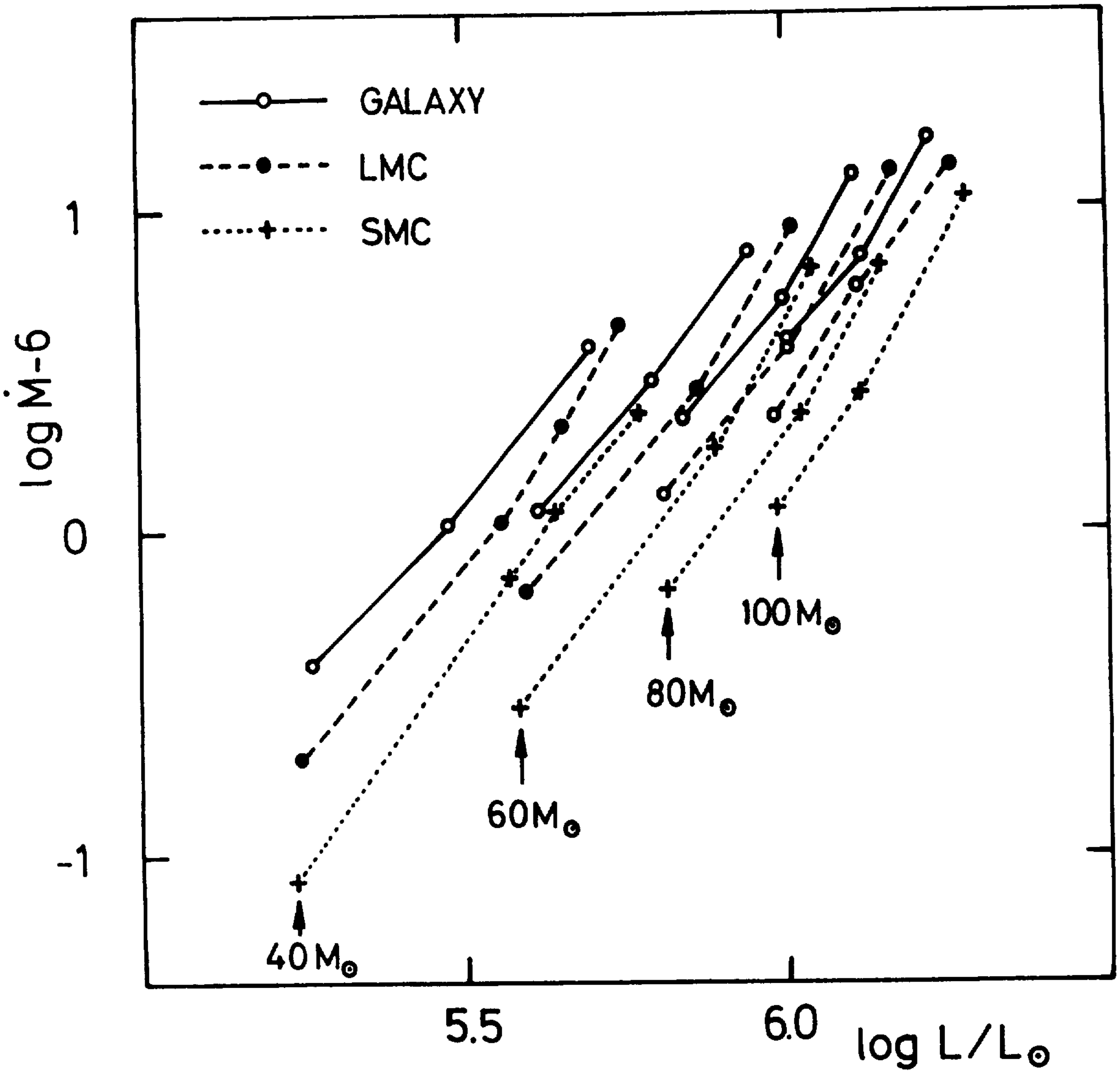


Figure 4-18a. Theoretical (log $\dot{M}$ /log L) relation along the same tracks as in Figure 4-16.

D. Validity of the Sobolev Approximation

The development of the radiation-driven wind theory in Section B was considerably simplified by the use of the Sobolev approximation $v_{th}/(dv/dr) << 1_{wind}$ (assumption c) in Sections B.1, B.2, and B.3). As a consequence of this assumption, no detailed radiative transfer was required to calculate the line forces, which were obtained simply by considering locally the velocity-induced line shifts $\Delta\nu_i$ toward the nonattenuated neighboring parts of the photospheric continuum.

This assumption has been subject to some criticism. It has been argued that the use of the Sobolev approximation throughout the atmosphere leads to inaccurate velocity fields and mass-loss rates (Weber, 1981; Leroy and Lafon, 1982). However, because the comparison between the Castor et al. (1975a) calculations and those by Weber and by Leroy and Lafon, who carried out exact radiative-transfer calculations

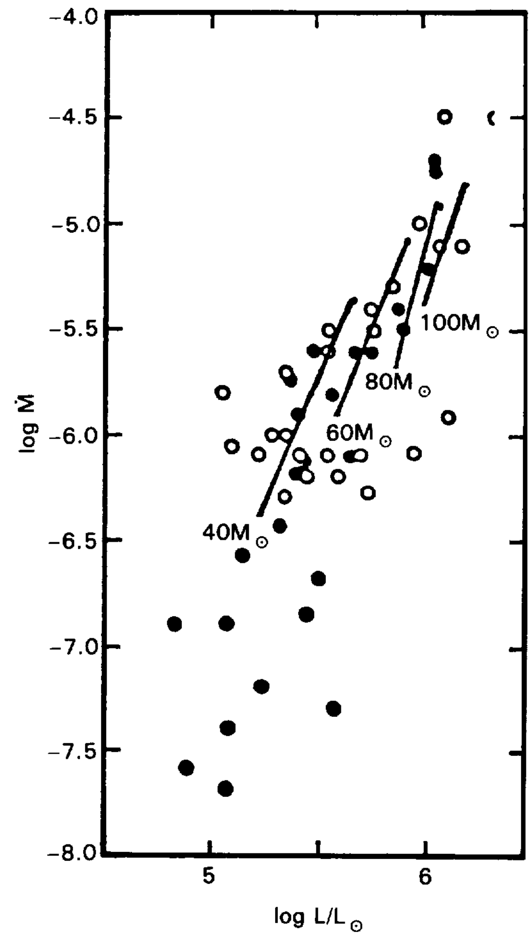


Figure 4-18b. Same as Figure 4-8, but including the calculations along galactic evolutionary tracks from Figure 4-18a.

instead of using the Sobolev approximation, were not made on a strictly differential basis, this conclusion needed further critical investigation.

Therefore, Pauldrach et al. (1986) investigated this problem in detail on a strictly differential basis. Parallel to the improved code described in Section B.3 using the Sobolev approximation "modified Castor et al."-theory (MCAK), they developed a completely independent complementary code in which the radiative force is calculated and the hydrodynamic equations are solved using a line-by-line solution of the exact (to order v/c) radiative-transfer equation in the comoving frame (CMF):

$$\mu \frac{\partial}{\partial r} I(\nu,\mu,r) + \frac{1-\mu^2}{r} \frac{\partial}{\partial \mu} I(\nu,\mu,r) - \frac{v_o}{c} \left((1-\mu^2) \frac{v}{r} + \mu^2 \frac{dv}{dr} \right) \frac{\partial I}{\partial \nu} = \eta(\nu,r) - \chi(\nu,r)\, I(\nu,\mu,r) \quad . \qquad (4\text{-}35)$$

Here $I(\nu, \mu, r)$,ν,7 (ν, r), $\chi(\nu, r)$ are the intensity, frequency, total emissivity, and opacity in the CMF. (For a detailed discussion of this equation, see Mihalas et al. (1975).) With H_ν defined as

$$H_\nu = \frac{1}{2} \int_{-1}^{1} d\mu\, I_\nu(\mu)\, \mu \quad , \qquad (4\text{-}36)$$

the radiative line acceleration by a single line i is then given by:

$$g^{i}_{rad} = \frac{4\pi}{c\,\rho\,(r)} \int \chi^{i}_{\nu}(r)\, H_\nu(r)\, d\nu \quad , \qquad (4\text{-}37)$$

where χ^i_ν is the line absorption as given in principle by Equation (4-16), but with a frequency now transformed into the CMF. The transformation between the CMF and the observer's frame reads (v/c << 1):

$$\nu_{CMF} = \nu_{obs} - \mu\, v_i/c \quad . \qquad (4\text{-}38)$$

For the representation of the line force, a sample of strong ($k_i/s_E = 10^6$), intermediate ($k_i/s_E = 10^4$), and weak ($k_i/s_E = 10^2$) lines has been chosen at three frequency regions longward of, shortward of, and at the flux maximum. After calculating the radiative forces for each of these lines by solving the CMF transfer equation, the individual forces were weighted according to the line-strength distribution function described by Equation (4-20) in Section B.2 so that, by definition, the total force is the same as the MCAK force in the outer parts of the wind. This procedure ensured a strictly differential study in which the acceleration laws were essentially identical well outside the sonic point, where the Sobolev approximation is certainly sufficient, but the line forces were allowed to differ around the sonic point, where the Sobolev approximation might be questionable and the correct CMF radiative transfer is required.

The result of those extensive calculations was that only small differences are introduced by the Sobolev approximation in the MCAK-theory. (For details, see Pauldrach et al., 1986.) This is demonstrated in Figure 4-19, a comparison of the wind-velocity fields, and in Table 4-5, which compares the results obtained by the MCAK-theory for the galactic O stars (Table 4-3) with those computed using the CMF force.

Although these results now appear to show that the approximations of the MCAK theory—Sobolev approximation, core-halo structure, no limb-darkening—are of little importance, we believe that it is necessary to continue with the more sophisticated CMF theory in the future for the following reasons. First, we cannot be completely sure that the agreement is simply caused by the renormalization procedure itself that fixes the CMF calculations in the outer layers. Second, we feel that, in the case of extended photospheres as for Wolf-Rayet stars (see Pauldrach el al., 1985) and extreme supergiants, the CMF calculations will be clearly more appropriate. Therefore, the calculation of self-consistent wind models using CMF and Abbott's full-line list will be part of the future work. This will also permit us to investigate in a realistic way the effects of multiple scattering on the velocity field. (See Section F.)

E. Superionization and Statistical Equilibrium in O-Star Winds

One of the striking peculiarities in the winds of hot luminous stars is the so-called "superionization," in which the presence of anomalously high stages of ionization like O VI and

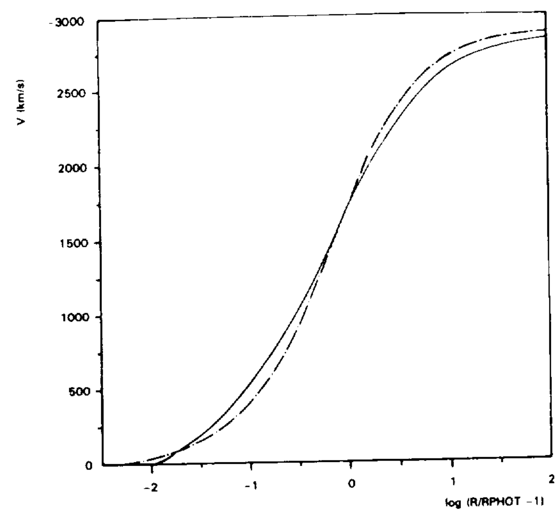


Figure 4-19. Theoretical velocity field for a typical O5f star obtained with the CMF line force (—) and the improved (MCAK) theory as described in Section B.3 (—·) (from Pauldrach et al., 1986).

Table 4-5

Computed Wind Parameters Obtained With CMF Stellar Wind Code Compared With Results Given in Table 4-3 Using the Improved (MCAK) Theory.

Star	$10^{-6}\ M_\odot$/yr $\dot{M}_{MCAK}$	$\dot{M}_{CMF}$	km/s v_∞^{MCAK}	v_∞^{CMF}
P Cyg	29.0	24.0	395	395
ϵ Ori	3.3	3.4	1950	1930
ζ Ori	1.9	1.7	2274	2140
9 Sgr	4.0	3.4	3480	3540
HD 48099	0.64	0.7	3540	3150
HD 42088	0.20	0.23	2600	2320
λ Cep	5.1	4.0	2500	2470

N V are meant. Two examples are given in Figures 4-20 and 4-21. The term "superionization" was chosen because the first non-LTE calculations for wind models with a temperature structure of $T_e \approx T_{eff}$ (cool winds)—as resulting from radiative equilibrium—were not able to reproduce these high ionization features. As a consequence, Lamers and Morton (1976) introduced the "warm-wind model," in which $T_e \approx 2 * 10^5$ K is adopted throughout the wind, and O VI is therefore produced by collisional ionization. Alternatively, the "thin-corona model" was suggested (Hearn, 1975; Olson, 1978; Cassinelli and Olson, 1979; Olson and Castor, 1981), in which a thin corona of $T_e \approx 10^6$ K at the bottom of the stellar wind emits X rays. These then produce O VI from O IV by Auger photoionization. However, both the warm-wind and the thin-corona models disagree with far-IR observations (Lamers et al., 1984). In addition, the observed X-ray spectra disagree with predictions of the thin-corona model (Cassinelli and Swank, 1983). (For a discussion of Waldron's (1984) results, see Pauldrach (1986).) Consequently, because both models could be ruled out, the problem of superionization remained unsolved.

A qualitative solution was suggested by the "shock model" (Lucy and White, 1980; Lucy, 1982b; Krolik and Raymond, 1985; Owocki and Rybicki, 1985). Here, instabilities in the wind produce randomly distributed shocks. In turn, these shocks can produce X rays with a spectrum close to the observations. It was therefore speculated as to whether these X rays are also present in the deeper layers of the wind where they produce O VI and N V, again by Auger ionization. However, this has not been worked out quantitatively, mainly because the details of the shock properties are still too uncertain. Simple, somewhat crude estimates (Pauldrach, 1986) lead to the conclusion that, in the inner part of the wind ($v < 0.9 v_\infty$), the high-ionization species may not be created by shock X rays.

This unhappy situation, in which the observed ionization stages cannot be reproduced by a self-consistent theory, is troublesome. (It

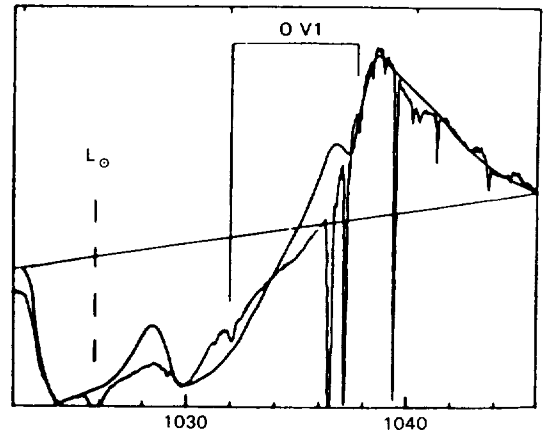


Figure 4-20. O VI P Cygni profile in the FUV spectrum of ζ Pup (O4f) and a computed profile based on a detailed "warm-wind model" (from Hamann, 1980).

is also the reason why mass-loss rates obtained from UV wind lines are regarded as extremely uncertain.) We have therefore decided to attack once more the ionization problem in hot-star stellar winds. We realized that all previous statistical equilibrium calculations in stellar winds used rather crude approximations—either ionization from the ground state followed by total recombination and electron cascades downward was adopted (Lamers and Morton, 1976; Hamann, 1980, 1981; Abbott and Lucy, 1985; etc.) or ground state ionization followed by direct recombination to the ground state only (Abbott, 1982a; Pauldrach et al., 1986; Kudritzki et al., 1986). The first approach assumes that all lines are optically thin (as in planetary nebulae or in H II regions); the second one assumes that all lines are optically thick (as in deep photospheric layers). In view of the densities (10^8 cm^{-3} $\leqslant n_E \leqslant 10^{12}$ cm^{-3}) in stellar winds, it is clear that neither of these assumptions can be correct.

Another crucial approximation concerns the ionizing radiation field. In almost all approaches, it has been represented by optically thin geometrically diluted photospheric radiation. For the most important ions, however, the ionization edges lie beyond the He II edge, where one can show that the radiation is

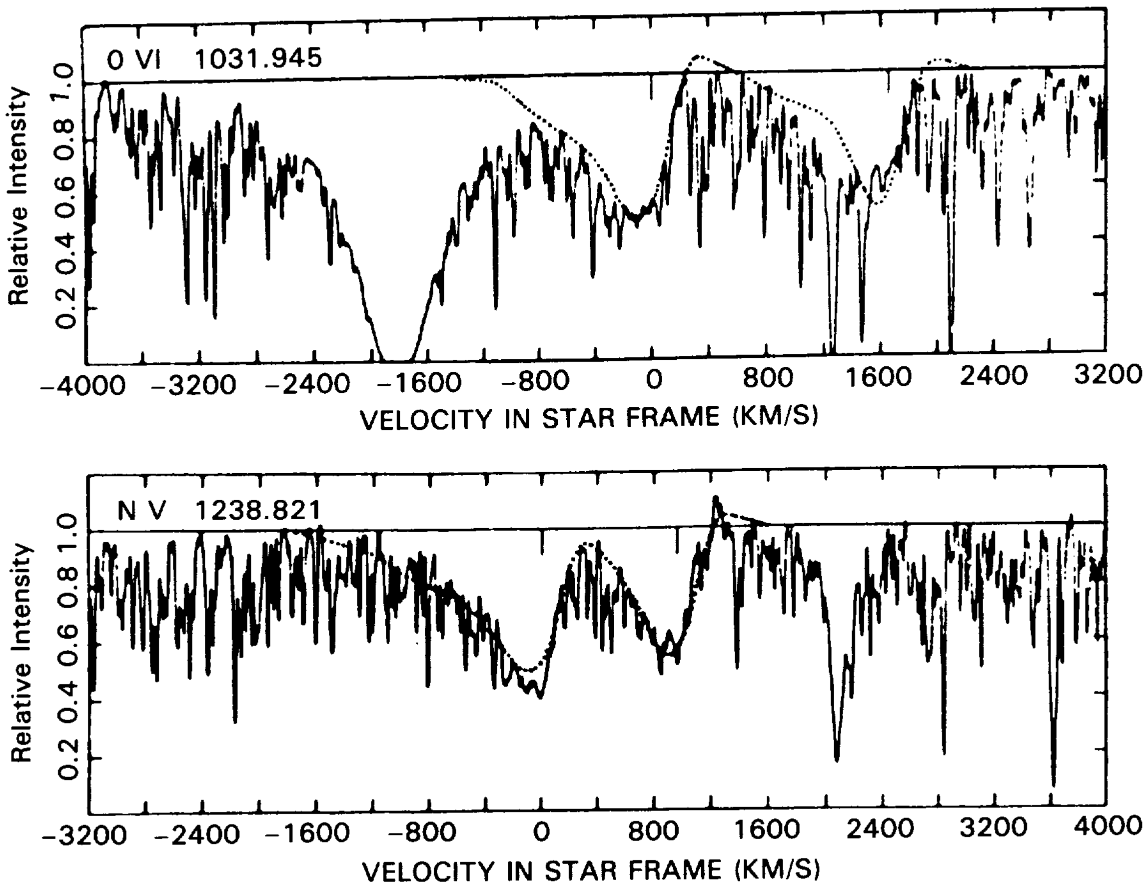


Figure 4-21. O VI and N V profiles with flat extended blue wings in τ Sco (O9.5 V). The dotted profiles are computed on the basis of a "warm-wind model" (from Hamann, 1981).

optically thick. Thus, in this region, the ionizing radiation is undiluted and is coupled to the local electron temperature.

We therefore decided to attack the ionization problem again from the viewpoint of the cool-wind model. The temperature structure was chosen ad hoc, following closely spherically extended non-LTE atmospheric models in radiative equilibrium, which take into account the wind-model density structure but only hydrogen and helium bound-free and free-free opacities (Gabler et al., 1987). Using $T_e(r)$, the full problem was treated self-consistently (i.e., we solved the full multilevel non-LTE rate equations, including electron collisions, for all relevant elements and ions simultaneously with the hydrodynamic equations of radiation-driven winds). Our calculations comprise in total: *133 ionization stages of 26 elements (H to Zn) with altogether 4000 levels and 10000 bound-bound transitions. Electron collisions are included and the correct continuous radiation field obtained from the solution of the spherical transfer equation is taken into account for bound-free transitions. The occupation numbers obtained in this way are then used to calculate the contribution of more than 100000 lines to the line force, using the line list of Abbott (1982a).* The new line force is then used again for the solution of the hydrodynamic equations, which yield new density and velocity fields. In turn, these allow the calculation of new occupation numbers. The process is iterated to convergence. (For details, see Pauldrach (1986).)

The atomic model for Ne IV is sketched in Figure 4-22 as one example for the treatment of the 133 ionization stages in this extensive work.

The first calculations have been carried out for ζ Pup because this object has well-studied wind properties and precisely determined stellar parameters (Kudritzki et al., 1983; Bohannan et al., 1986): T_{eff} = 42000 K, log g = 3.5, $R/R_{\odot}$ = 19.

We report the following results of these first calculations:

(a) For frequencies beyond the He II edge, the ionizing radiation is indeed optically

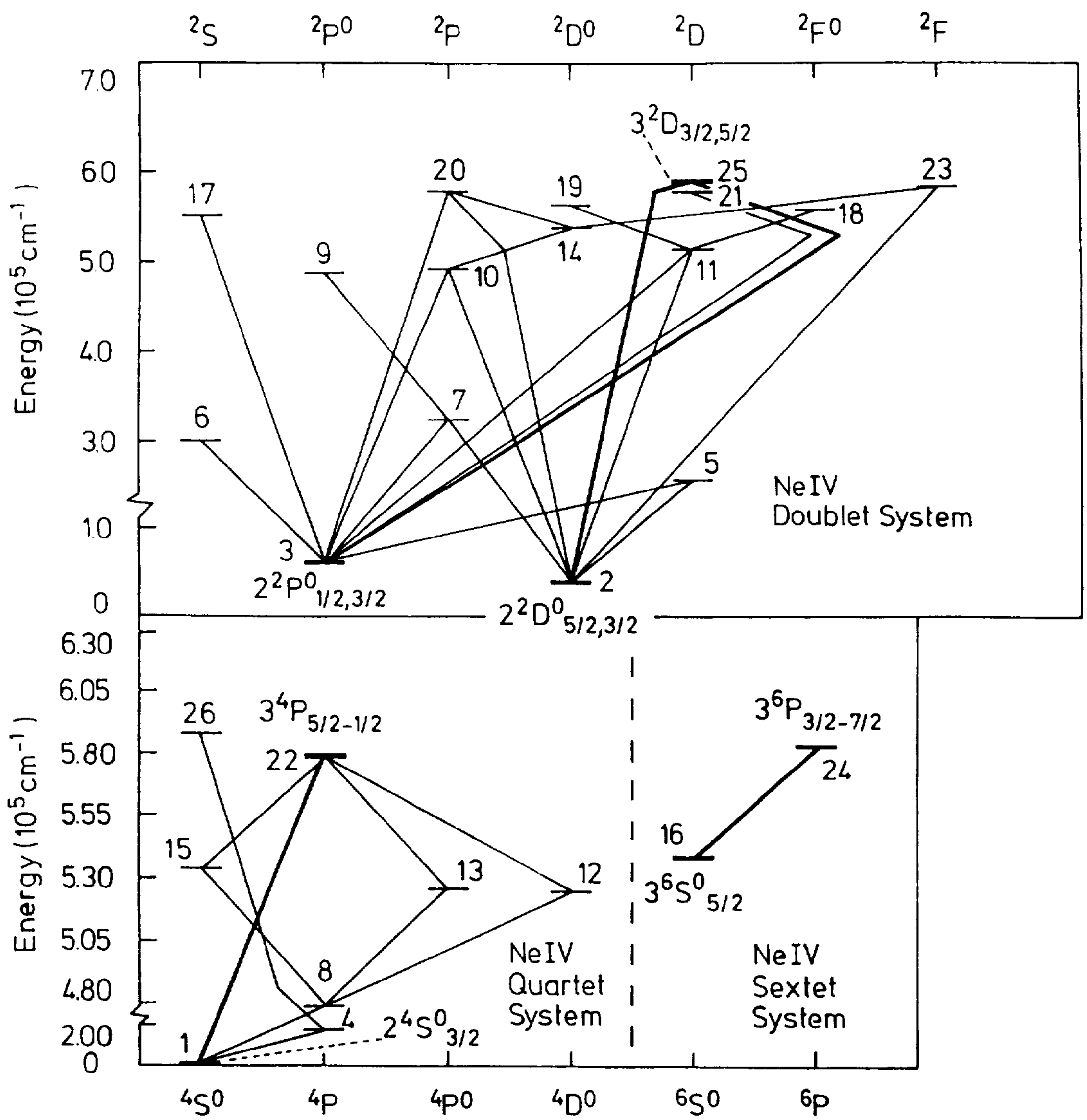


Figure 4-22. Grotrian diagram of Ne IV with all radiative bound-bound transitions taken into account as an example for the treatment of the 133 different ionization stages in the detailed multilevel statistical equilibrium calculations (from Pauldrach, 1986).

thick up to several stellar radii (Figure 4-23). An enormous O V absorption edge occurs in the wind.

(b) A dramatic shift to higher ionization stages results from the solution of the full non-LTE multilevel equations (Figure 4-24). The ground levels of the higher ionization stages are now orders of magnitudes more strongly populated than in the older more approximate calculations. The effect is strongest in the deeper layers, but remains at high velocities as well. The main mechanism that produces this effect in the inner part of the wind is sketched in Figure 4-25. Collisional excitation of low-lying levels with ionization wavelength $\lambda_{Ion} > 228$ Å is followed by photoionization caused by the powerful photospheric radiation at wavelengths longward of the He II edge. A second mechanism, which works mainly in the outer wind regions, is the increased direct radiative ionization from the ground state due to the optically thick undiluted local radiation field with $\lambda > 228$ Å.

(c) The shift to high ionization stages also has a dramatic influence on the O VI ionization (Figure 4-26). While the old approximations fail by more than 6 dex, the new calculations agree very well with the observations. Figure 4-27 demonstrates that, as a result, clearly observable O VI is produced by the calculations. *Thus, the longstanding problem of "superionization" can be solved for cool winds without any other source of ionization.*

We stress, however, that $T_e(r)$ is still chosen ad hoc (as described previously), neglecting line opacities (hydrogen, helium, and metals), in the radiative equilibrium. Thus, nothing can be said at present about the importance of nonradiative heating. Although the wind temperature is now much lower than previously discussed in connection with "superionization," it cannot be excluded that additional nonradiative heating is needed, if future radiative equilibrium models including metal line-cooling should predict too low temperatures compared with the observations. Despite this open question, we regard the present results to be a decisive step in our understanding of the ionization mechanisms in hot-star winds.

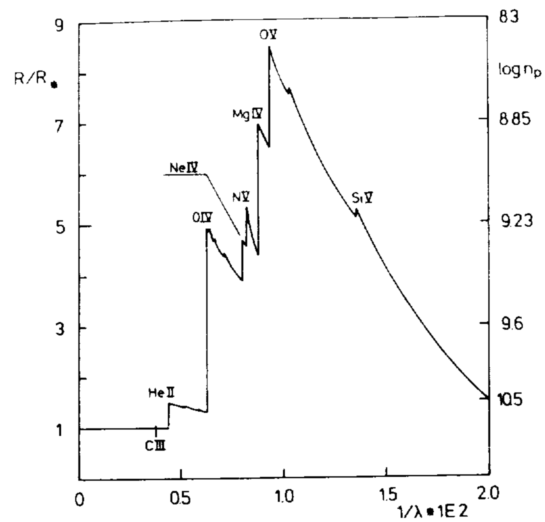


Figure 4-23. Radius of optical depth equal unity as a function of reciprocal wavelength in the wind of ζ Pup (from Pauldrach, 1986).

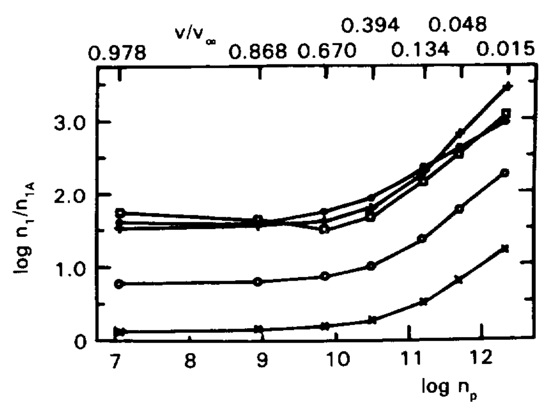


Figure 4-24. Logarithm of the ground state occupation number in the full multilevel non-LTE divided by the value obtained in the old approximation (cf. Abbott, 1982a): ● = N V, O = S VI, × = Ne IV, + = A VI, and □ = Fe V. The depth parameter is the logarithm of proton density (bottom) or the velocity in units of v_∞ (top) (from Pauldrach, 1986).

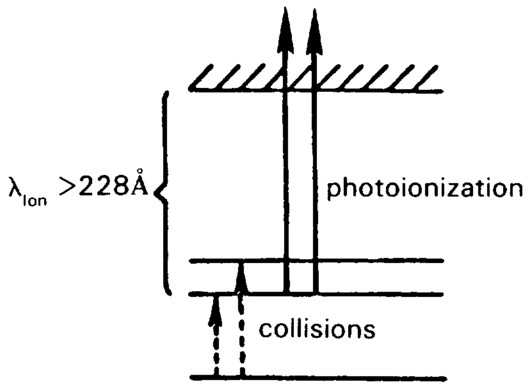


Figure 4-25. Sketch of the ionization mechanism of highly ionized species in the stellar wind.

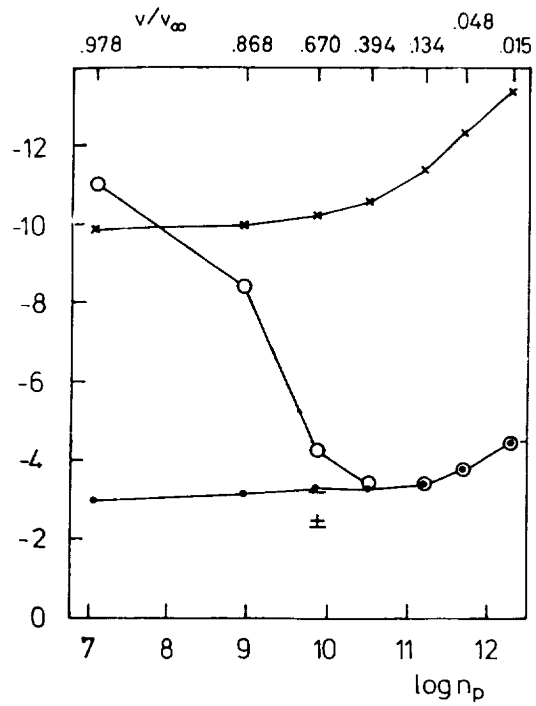


Figure 4-26. Logarithm of O VI ionization fraction versus same depth scale as in Figure 4-22: × = old approximation (cf. Abbott, 1982a), • = full non-LTE multilevel calculations with $T_e = T_{eff}$ = const., o = full non-LTE, but with slow-falling $T_e(r)$ according to the spherical extended radiative equilibrium code by Gabler et al. (1986), + = observed value by Lamers and Morton (1976), and - observed value by Hamann (1980).

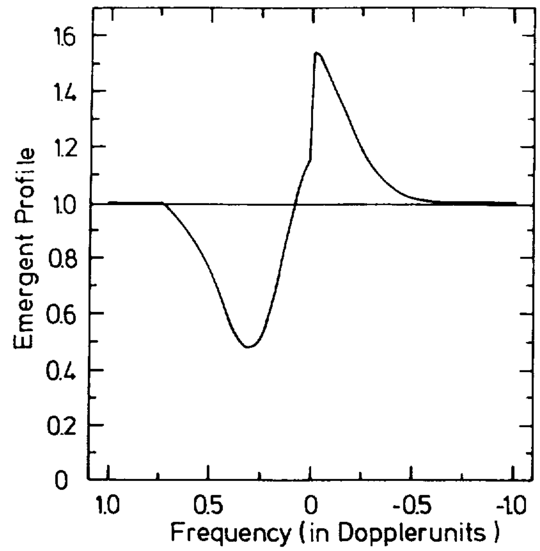


Figure 4-27. Calculated red doublet component of the O VI resonance line using the ζ Pup wind model with slow-falling radiative equilibrium temperature.

(d) The back-reaction of the improved occupation numbers on the wind dynamics is also significant. Excellent agreement with the observed wind parameters is obtained:

$\dot{M}_{obs}$	$\dot{M}_{calc}$	v_∞^{obs}	v_∞^{calc}
$10^{-6}\,M_\odot$/yr	$10^{-6}\,M_\odot$/yr	km/s	km/s
4–6	4.6	2450	2200
		2660	

The few hundred km/s missing in v_∞ are, in our opinion, related to the effects of line overlap and multiple scattering.

(e) Far-IR observations (IRAS) give information about density and stellar-wind stratification close to the photospheric surface (Lamers et al., 1984). It is therefore interesting to compare the prediction of the self-consistent wind model with the observations. This is done in Figure 4-28, which shows that this part of the spectrum is now described properly by the theory as well.

F. Future Improvements of the Theory

From the foregoing section, we conclude that, for luminous OB stars, the concept of radiation-driven winds is the right way to describe quantitatively the rapidly expanding envelopes of hot stars. Although the work done thus far comprises the effort of many people over many years, it is only a first step in this direction. Clearly, the next step will be to include the effects of spectrally neighboring lines. The velocity-induced Doppler shifts lead to overlaps of line transitions so that a photon emitted in the wind by one line can be absorbed again by another red-shifted line. (Figure 4-29 gives a typical example for this situation.) This "multiple scattering" in different line transitions can have two important effects:

(a) Modification of the occupation numbers in the wind: From Section E, it is evident that a detailed non-LTE multilevel treatment is required for the proper calculation of the occupation numbers of the individual ions. These occupation numbers are partially determined by the radiation field in the line transitions, which is changed because of the line overlap. Until now, the latter has been neglected.

(b) Modification of the stellar-wind dynamics: The multiple absorption in different line transitions will lead to a multiple transfer of photon momentum to the stellar-wind plasma, which will affect the radiative force and, consequently, the wind dynamics.

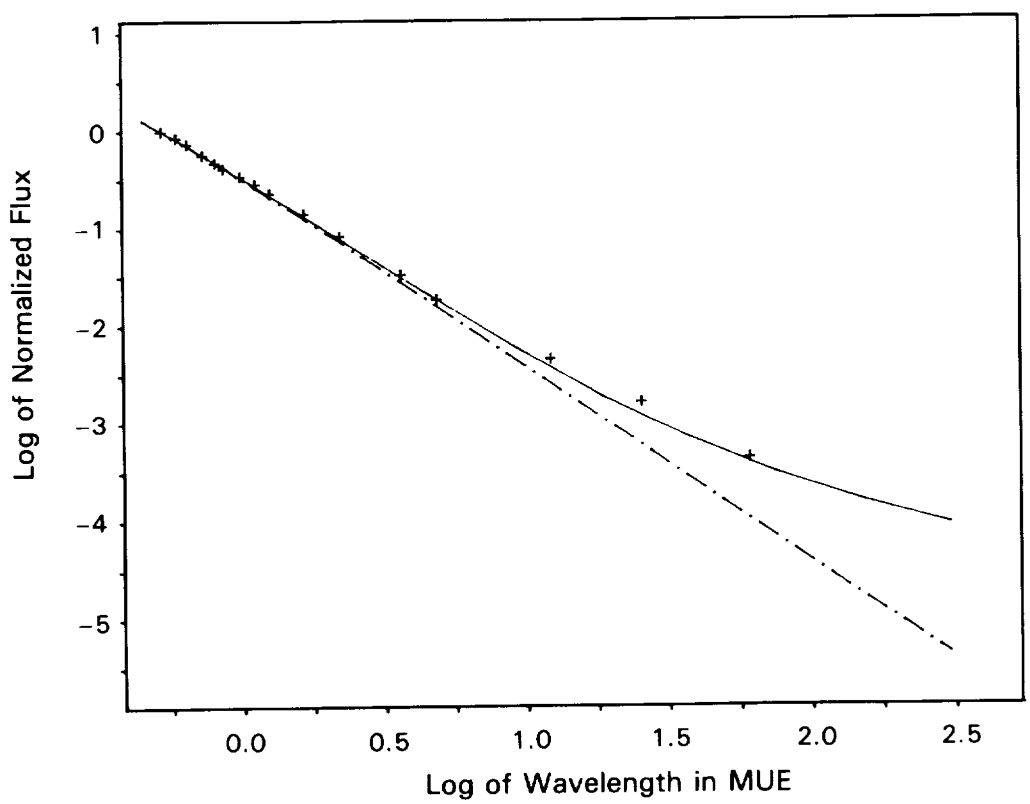


Figure 4-28. Observed IR energy distribution (+) as given by Lamers et al. (1984) compared with a spherical non-LTE model with stellar wind (—) from Gabler et al. (1986) and a plane parallel model without wind (—·) from Kudritzki et al. (1983).

A variety of attempts have already been undertaken to investigate this effect. Panagia and Macchetto (1982a) used a simple model to study the influence on the velocity field, but neglected the back-reaction on mass-loss rate and occupation numbers. Abbott and Lucy (1985) used Abbott's (1982a) line list and adopted a fixed velocity field v(*r*) and a Monte Carlo technique for the multiline radiative transfer to calculate the momentum transfer from the continuous photospheric radiation to the stellar-wind plasma. By this method, they were able to compute a mass-loss rate for ζ Pup that agrees well with the observed value. Because the velocity field was a priori adopted and kept fixed, however, nothing could be said with respect to back-reactions on the velocity field. In addition, the influence on the occupation numbers, calculated with the approximate non-LTE discussed in Section E, could not be considered at this stage. Friend and Castor (1983) recalculated one of the models by Castor et al. (1975a), taking into account multiple scattering in a self-consistent but simplified statistical way. Much larger values for v_∞ were found (the individual values depending on the parameterization of the line force), but from our point of view, it was not possible to clearly disentangle the influence of the finite cone-angle correction factor (Section B.3) and the influence of multiple scattering in this work. In any case, the importance of multiple scattering for self-consistent models was clearly indicated by this paper (and the two others just cited), so that a more realistic and refined treatment of this effect appears to be inevitable.

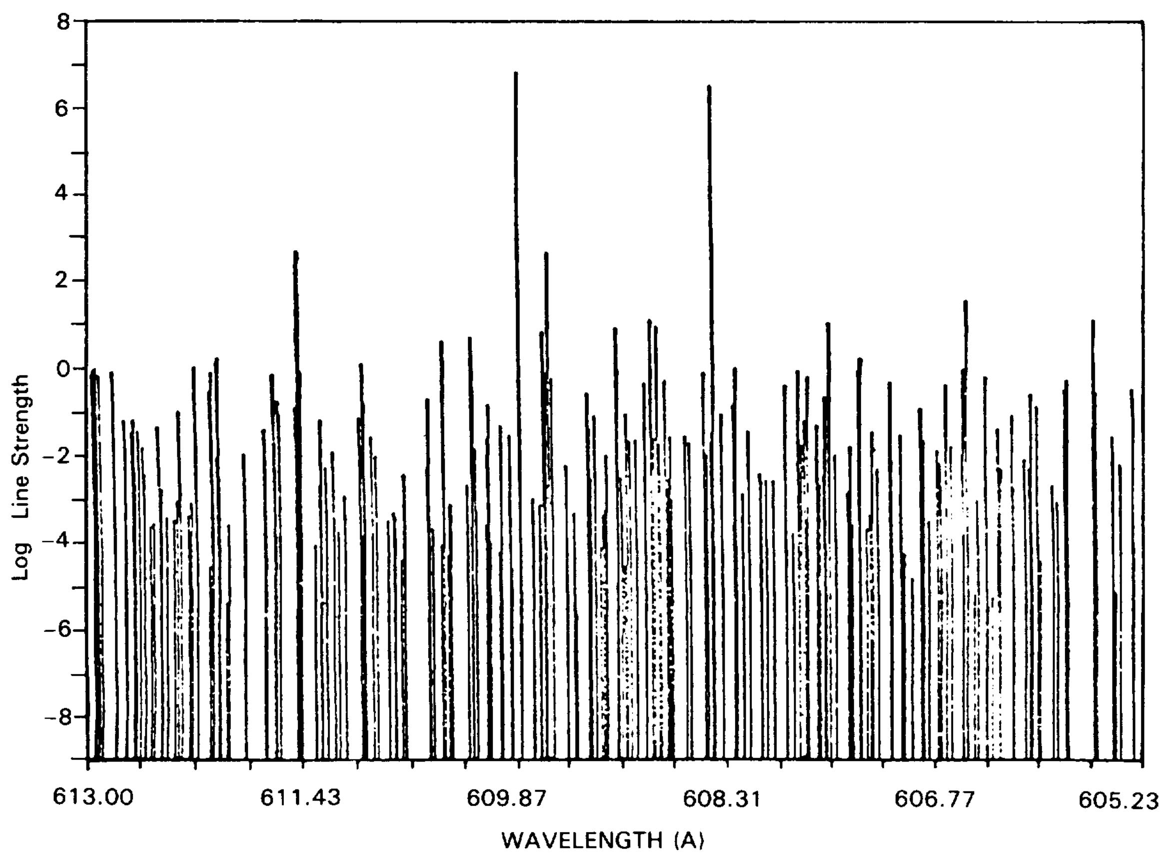


Figure 4-29. Distribution of line strength (k_i/s_E) in a typical O-star EUV spectral interval of width $2\Delta\lambda_\infty = 2\lambda\, v_\infty/c$. In a stellar wind, photons emitted by the bluest lines can, in principle, be reabsorbed by all redward lines in this interval, which contains 244 lines (computed by J. Puls).

In our opinion, the only way to test the effect quantitatively is to extend the calculations described in Section E, based on a realistic line list and detailed multilevel non-LTE calculations, so that multiple scattering is included. At first glance, this appears to be a hopeless task because all the multilevel rate equations are directly radiatively coupled due to the overlapping line transitions. However, the treatment of the problem within our group is well advanced, and first results will be forthcoming. With this important effect included, detailed stellar-wind model calculations will be undertaken not only for OB stars but also for Wolf-Rayet stars. For the latter case, the results of Pauldrach et al. (1985) indicate that the situation is not completely hopeless.

Another crucial topic is the self-consistent treatment of the energy balance in the stellar wind. A first step will be to calculate radiative equilibrium models (see Section E), including metal line absorption in the energy balance. In addition, however, it will be extremely important to investigate how the radiative equilibrium is affected by additional dissipation of mechanical energy. Besides some test calculations introducing the ad hoc dissipation of a certain fraction of mechanical energy, we suggest the use of the detailed and improved wind models with realistic line forces for a refined analysis of the stability of the acceleration mechanism and the wind flow. (For an outline of the basic methods, see Owocki and Rybicki (1985).) The results can then be taken for the investigation of the properties of possible acoustic waves or shocks, particularly the EUV and X-ray radiation field emitted by the latter (see Krolik and Raymond, 1985; Lucy, 1982b, 1986). This will be extremely important for the reliability of the ionization calculations.

Another point of interest is the variability of stellar winds. Until now, only a stationary theory has been developed, which is unable to make clear predictions with respect to variability. On the other hand, the variability in stellar-wind lines is clearly observed. (See the contributions by D. Baade and H. Henrichs in this volume.) In view of this, the following mechanism looks very interesting. Mass-loss rates and velocity field are very sensitive to the stellar parameters, as our calculations have shown. This means that even small pulsations must be reflected in the wind. For β Cep pulsators (σ Sco and BW Vul), we have checked this in our stationary models (Pauldrach, 1985) and found clear effects. Of course, in view of the short periods, time-dependent calculations will be necessary to investigate the effects of pulsation on the wind. It is interesting to speculate as to whether the observed "narrow line components" can be produced in this way or whether additional turbulence or nonmonotonic behavior of the velocity field (Hamann, 1980, 1981; Lucy, 1983, 1984a, 1986), which appears to be observed, can be induced. It might also be a mechanism for additional shock production and energy dissipation in the wind.

G. Observational Tests of the Theory

Until now, the observational tests of the theory consisted mainly of comparisons between observed and computed mass-loss rates for samples of luminous OB stars with more or less well-defined stellar parameters. Because we know that both the observed wind properties inferred from the plasma diagnostics of stellar-wind features and the calculated properties depend crucially on the adopted stellar parameters of luminosity, mass, temperature, radius, and *abundances*, precise determination of these parameters is necessary for a real test of the theory. Fortunately, the progress in the theory of hot-star photospheres and the dramatic improvement of observational spectroscopic techniques in recent years now allows a determination of stellar parameters with high precision even for objects as faint as O main-sequence stars in the Magellanic Clouds. (For a recent review, see Kudritzki (1985) and Kudritzki and Hummer (1986).) Therefore, we suggest that the improved theory of radiation-driven winds may be tested in more detail. Non-LTE analyses of photospheric spectra of individual galactic and LMC/SMC objects will have to be carried out to obtain $T_{\rm eff}$, $\log g$, L,

$\dot{M}$, and *abundances*. These parameters will permit the calculation of wind models that can then be compared in detail with the observations. We stress the importance of abundance determinations in this context. Since the contamination of photospheric material with CNO nuclear processed material (as a result of stellar evolution in the presence of mass loss) is to be expected theoretically and is also clearly observed in some cases (Kudritzki et al., 1983; Butler and Simon, 1985; Bohannan et al., 1986; Schonberner et al., 1986; Kudritzki and Hummer, 1986), the usual assumption of solar abundances for galactic objects is not of high reliability, particularly if the UV C IV or N V lines are used for the observational stellar-wind investigations. For Magellanic Cloud objects, the situation is even worse because our information about massive-star abundances in this case is rather crude and indirect. Consequently, a careful test of the reliability of the stellar-wind theory must be carried out simultaneously with a careful determination of stellar parameters, including abundances.

It will be very important to study in detail once more those objects that permit accurate observations in a wide spectral range, including the radio and the far-IR domain of the spectrum (ζ Pup, P Cyg, etc.) on the basis of the improved theory. For these objects, it should be possible to test whether the theory predicts (on the time-independent average) the correct density, velocity, and ionization stratification. This will also make it possible to investigate whether the full non-LTE calculations described in Section E yield reliable ionization fractions in the wind and therefore permit mass-loss rates from UV wind lines to be determined with much higher precision than before. If so, this will be a breakthrough in UV stellar-wind diagnostics.

A striking observational feature of hot-star winds is the luminosity-dependence of the Si IV resonance lines that became evident from the work by Walborn et al. (1985), who systematically studied the morphology of high-resolution IUE spectra. It is obvious that a reliable wind theory must be able to reproduce such general effects. The same holds for the disappearance of "superionization" features at lower effective temperatures around spectral type B0 (see, for instance, Abbott et al., 1982).

Once the theory is able to include the effects of multiple scattering in a realistic way (see Section F), wind models for Wolf-Rayet stars will have to be constructed in order to investigate whether the theory also works in this case or whether another mechanism is needed. In our opinion, this is an open question (see Pauldrach et al., 1985) that can be decided only after the stellar parameters have been determined with some precision. For this purpose, spectroscopic methods as developed by Hillier (1983, 1984), Hamann (1985a, 1985b, 1986), and Hamann and Schmutz (1986) are extremely important.

After the enormously fruitful period of 8 years of the IUE satellite, it is important to remember that almost all of our information about the crucial stellar-wind lines in the FUV (912 Å $< \lambda <$ 1150 Å) dates back to observations taken with the Copernicus satellite more than 10 years ago. Although it might appear too trivial to be pointed out in a book such as this, we wish to make it clear that, for O stars, the FUV and the EUV are the decisive spectral regions in which most of the intrinsic stellar flux is emitted. Consequently, future observations using the highly efficient UV spectroscopic technology of today will put the most severe constraints on the theory. The same will hold for the spectroscopic X-ray missions planned for the next decade. With spectra of sufficient resolution in the soft X-ray domain, we will be able to determine how the stellar winds are able to produce this radiation. Besides the powerful tools of plasma diagnostics of X-ray emission lines, the simple Doppler shift of these lines will give us information about the region in the wind (inner and/or outer layers) from which this radiation originates. Although all this is a long way off, we have no doubt that the astrophysical importance of mass loss will clearly justify the effort.

H. Acknowledgments

It is a pleasure to thank our colleagues at the Universitäts-Sternwarte Munchen for the stimulating interest in our work and the many helpful discussions. This includes in particular Dr. D.G. Hummer, who as a Humboldt-foundation awardee, spent 1 year with us. We also wish to thank Dr. D.C. Abbott for his useful comments on our work and for providing us with his line list. In addition, we are grateful to Drs. R.N. Thomas and V. Doazan for their frank and constructive criticism of our work, which helped to improve the manuscript. Special thanks to Dr. Keith Butler for the careful reading of the manuscript and to Klaudia Killen for the engaged preparation of the manuscript.

V. INTRINSIC VARIABILITY IN ULTRAVIOLET SPECTRA OF EARLY-TYPE STARS: THE DISCRETE ABSORPTION LINES*

A. Introduction

The long missions of the Copernicus and IUE satellites made it possible to monitor ultraviolet spectra of early-type stars on time scales from less than 1 hour to many years. Intrinsic variability (i.e., not due to binary effects) is frequently found in the resonance lines of O VI (1032, 1038 Å), Si III (1207 Å), N V (1239, 1243 Å), Si IV (1394, 1403 Å), C IV (1548, 1551 Å), Al III (1855, 1863 Å), the nonresonance) line of C III (1176 Å), etc. (i.e., spectral lines that indicate the presence of a stellar wind). Like most stars across the HR diagram, early-type stars are thought to lose mass in the form of a stellar wind. Determining mass loss on the basis of spectral line shapes is therefore one of the main tasks of the ultraviolet spectroscopists. (Of course, other methods are also available.) To assist in this task, theoreticians have been developing models for stellar winds and their structure. This enables a prediction to be made for a specific spectral line shape, depending on the stellar-wind structure parameters such as mass-loss rate, velocity, density and ionization structure, presence of X rays, etc. A comparison of predicted and observed line profiles then gives numerical values for the desired parameters.

In theoretical models, however, the basic assumption is usually made that the stellar-wind structure is time-independent. (For a recent review, see, for example, Abbott, 1985.) On the other hand, it has become obvious from the steadily accumulating UV data that the majority of spectral lines that are stellar-wind indicators are, to some level, variable on all time scales—in some cases even as short as ½ hour (i.e., comparable to a typical flow time scale of a stellar wind (defined as the stellar radius divided by the observed radial velocity)). Of particular concern is the interpretation of the (generally) blue-shifted "discrete absorption components." These are also frequently called narrow components, narrow lines, shifted narrow components, etc. (*narrow* to emphasize the contrast to the often-observed broad underlying P Cygni profile). In fact, the best description is "unexpected" absorption. In most cases, the components are narrow and/or shifted, but not always. This is why we refer to the phenomenon in general as "discrete" absorption components. These spectral features are beyond any doubt identified as mostly resonance transitions of abundant ions that are indicators of mass outflow. The most well-known example is given by Morton (1976), who observed the N V doublet in ζ Oph, O9.5 V(e), with the Copernicus satellite (Figure 4-30). Both members of this doublet are found at their correct relative position, as well as relative strength, but are observed to be blue-shifted corresponding to 1400 km/s, whereas the width is only 200 km/s. The earliest UV spectrum of ζ Oph (recorded around 1968 by Smith, 1972) shows essentially the same features. Other early examples may be found in Underhill (1975).

*Huib Henrichs, Joint Institute for Laboratory Astrophysics, University of Colorado and National Bureau of Standards, Boulder, Colorado 80309-0440, and Astronomical Institute, University of Amsterdam, Roetersstraat 15, 1018 WB Amsterdam, the Netherlands.

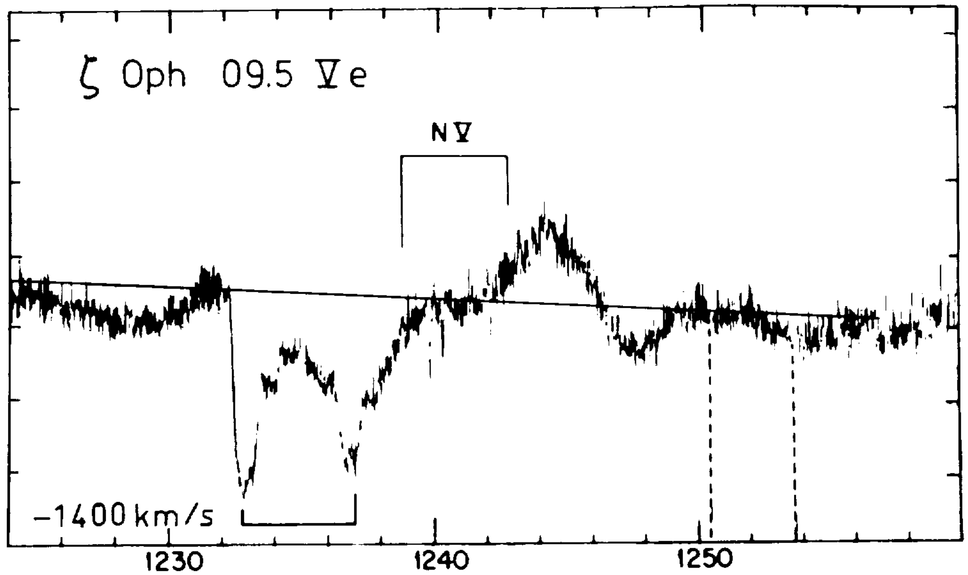


Figure 4-30. Copernicus spectrum of ζ Oph (Morton, 1976), showing shortward displaced narrow absorption lines of the resonance doublet of N V.

In a survey to detect mass outflow, Snow and Morton (1976) discovered discrete absorptions in many other stars (ζ Pup, O4 I(n)f; 15 Mon, O7 V((f)); λ Ori, O8 III((f)); ρ Leo, B1 Iab; ξ Per. O7.5 III(n)((f)); etc.). These stars appear to be superimposed on the blue absorption wing of a P Cygni profile. The first systematic investigation of their presence (i.e., without studying variability) was presented by Lamers et al. (1982), who found 17 stars with high-velocity discrete absorptions in a sample of 26 OB (non-Be) stars, which indicated that this phenomenon is rather common in such stars. Lamers et al. noticed an important property in their sample: if the discrete components in a given spectrum are found in different ions, they always show the same velocity, and this velocity is in all cases less than the terminal velocity of the wind (being identified with the steep blue edge of saturated profiles of other P Cygni lines when available). (See Figure 4-34c for a clear example.)

The discrete absorption lines have also been discovered in a number of Be stars (γ Cas, B0.5 IVe (Hammerschlag-Hensberge, 1979); 59 Cyg, B1 Ve (Doazan et al., 1980); ω Ori, B2 IIIe (Peters, 1982b); etc.) as well as in some subdwarfs (e.g., HD 128220B sdO (Hamann et al., 1981; Bruhweiler and Dean, 1983). In the case of Be stars, it is well known that the P Cygni lines are not well developed (in the sense that the emission is absent), and the discrete components are superposed on the asymmetric absorption wings. Multiple absorption components are frequently found with different velocity and strength (e.g., 59 Cyg, γ Cas, ω Ori, 66 Oph, and HD 128220B). A detailed summary of observed properties is given in Section C.

Connected with the presence of the discrete components is the issue of their *variability*. Typical examples are collected in Figure 4-34. (See also Figure 4-44.) As mentioned previously, the observed time scales can be as short as a ½ hour. In some cases, particularly the less luminous Be stars (i.e., with later spectral type), the entire absorption profile is variable and may sometimes disappear (e.g., θ CrB, B6 IIIe (Doazan et al., 1984)). (See also Figure 4-34g.) In this case, we need two time scales to describe the variability: (1) the short time scale which characterizes the variations in a given discrete component, and (2) a much longer time scale which characterizes the activity level, where "active" indicates the presence of the (variable) discrete components.

Quite a few explanations for the *origin* of these lines, which are definitively formed somewhere in the expanding stellar wind, have been proposed during the past 10 years. None of these explanations appears to be accepted by the different investigators. The main concern is to explain the observed variability of the discrete absorptions. The pioneering work by Snow (1977) contains the first systematic study of the variable character of UV spectra of early-type stars. Snow found in many of his 17 stars conspicuous changes over a few years, mostly in strength of the high-velocity components, although some changes in velocity were also observed. His sample contained O I-V stars and a few B supergiants. Many reports on variable UV lines in OB stars have appeared. (See for example, Snow (1979), Underhill and Doazan (1982), Marlborough (1982), and Henrichs (1984).) A more recent compilation appears in Section D.

The first systematic investigation of the discrete high-velocity absorption lines *in one particular star* was made by Henrichs et al. (1980, 1982, 1983) for γ Cas, B0.5 IVe. Drastic changes both in strength (Figure 4-31) and in velocity of narrow features were observed in the majority of 28 high-resolution IUE spectra. The time scales involved varied between 1 week and 1 month. A remarkable "memory" in the star, γ Cas, has been found in the sense that, when the discrete absorption lines were strong, they were found at a systematically lower velocity than when they were weak, even though they sometimes completely disappeared between such different episodes. (This "memory" of this star did not appear to hold over more than 5 years (Doazan, private communication).) Henrichs et al. (1983) proposed that the time scale involved for the appearance of the narrow absorption lines might, in fact, be of the order of hours. In addition, they gave arguments that this behavior might not be unique for this single Be star, but in fact, would be representative for all early-type stars. An actual recording of such an "appearance" event

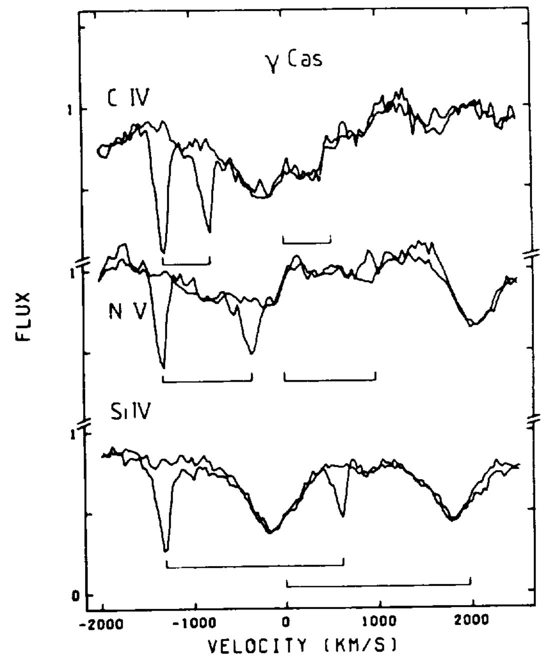


Figure 4-31. The striking difference between two spectra with and without discrete absorptions in γ Cas (Henrichs et al., 1983). In each row, the velocity scale is relative to the rest wavelength of the principal line of the doublet. Both undisplaced and displaced doublet positions are indicated.

has been looked for and observed in ξ Per, O7.5 III(n)((f)). (See Section D.) Since then, more systematic surveys concentrating on multiple spectra of a number of stars have been carried out. The most important survey study is the paper by Prinja and Howarth (1986), which contains a detailed analysis of available ultraviolet IUE spectra of more than 20 OB stars and draws systematic conclusions. (See Section C.)

Although this review is supposed to be concerned mainly with O stars, it will be obvious to the reader that the nature of the UV variability is such that we must inevitably include information about B and Be stars.

Motivated by these considerations, we devote the present review to the following topics:

(a) *Statistics:* In what fraction of early-type stars are discrete UV absorptions found and is their occurrence possibly related to spectral type, luminosity, rotation, Be characteristics, etc.? It will appear that, generally speaking, more than 60 percent of all OB stars show discrete absorptions at high velocity, whereas there is strong evidence that the discrete absorptions are, in fact, always variable. The fraction must be considered as a conservative lower limit. Another major conclusion is that, among the less luminous B stars, only Be stars show discrete components, in contrast to "normal" B stars, in which not much (if any) variability has been seen.

(b) *Variability:* A wide range in time scales of observed variations in both velocity and strength has been observed. We shall present significant examples of such variability on time scales as short as ½ hour, comparable to a typical flow time scale of stellar winds. A second, much longer, time scale is needed to describe the presence of discrete components in Be stars.

(c) *Models:* A short description of current models will be given. At present, none of them is capable of explaining all the observed properties of the variable discrete absorption components. However, some models explain reasonably well the general behavior of these lines in a few particular stars. No clear conclusion can be drawn about the origin of this phenomenon, although interesting suggestions have been made.

(d) *Related Observations of Interest:* High signal-to-noise optical spectroscopy has revealed that many early-type stars are in fact nonradial pulsators with varying amplitude and/or, to a lesser extent, modes. In a few cases, a correlation between variability in pulsation behavior and UV line profiles has been observed. This is a new and promising area in which only a first step has been made.

In summary, at present, it is at least the author's belief that the understanding of the line-profile variability is crucial for understanding the structure of the stellar winds of early-type stars generally, as well as for comprehending the unexplained Be phenomenon, a viewpoint which was a few years ago perhaps not as widely shared. The reader is referred to the book on B and Be stars (Underhill and Doazan, 1982) for earlier views regarding this matter.

Any realistic model describing stellar-wind structure must account for the observed time-dependent character. It is well recognized that stability analyses of stellar winds driven by radiation pressure, which is generally accepted to be the major driving mechanism in early-type stars, have shown that such flows are unstable against several kinds of perturbations. (See also Section E.) None of these models, however, can predict or even describe the behavior of the discrete absorptions as a function of time. For this reason, the major emphasis of this review is on the observational facts.

B. Statistics

The first question which must be addressed is: How common is the presence of discrete absorption components in the UV spectra of early-type stars? We shall see that they occur in all kinds of early-type stars, not specifically in "peculiar" stars, and that the observational evidence indicates that in fact encountering discrete absorption components is the rule rather than the exception. Another evident question is whether known physical properties of stars might be related to the occurrence of discrete absorptions.

Profiles of resonance lines in the UV spectra of many early-type stars can be found in the literature, much of which was published for the purpose of determining mass-loss rates and for showing variability. Major systematic surveys for the presence of narrow absorption lines are presented by Snow and Morton (1976, 47 stars), Snow (1977, 17 stars), Lamers et al. (1982, 26 stars) and Abbott et al. (1982, 53 stars), all of which are based on Copernicus spectra. Lamers et al. (1982) found discrete absorptions in 17 of 26 OB (non-Be) stars. Garmany (1982) collected data on 45 luminous stars. The most detailed analysis has been made by Prinja and Howarth (1986) on 21 stars (again, no Be stars).

Recently Henrichs and Wakker (1987) analyzed published data of both the Copernicus and the IUE satellites (including many spectra from the IUE data bank). The resulting sample of 241 stars was examined for the presence of discrete absorption lines, and possible ways that properties of these lines might depend on a number of stellar parameters were investigated. Grady et al. (1987b) studied 62 Be stars and 45 normal B stars. Results of these studies will be summarized later. It should be stressed that detailed studies of individual stars indicate that the features are always variable.

1. Distribution in the HR Diagram. We summarize here the main results of Henrichs and Wakker (1987), whose work incorporates earlier results. Their sample consists of stars of spectral type O3 to B7 with luminosity class I to V, including Of, Oe, Be, and a few subdwarf stars. Most of the information exists on the N V, Si IV, and C IV doublets. A star was adopted in the survey only when at least two of these doublets had been inspected. In view of the transient character of the discrete absorption lines, all available spectra were scanned for their presence. A common property of occurring at the same velocity in different ions aided the identification. If the presence of discrete absorptions could be established with confidence in at least one spectrum, the star was classified as "yes," otherwise as "no"; 16 stars (of 241) were discarded because their resonance lines were heavily saturated P Cygni profiles, making identification impossible. Table 4-6 gives a number of representative examples of stars in which (variable) discrete absorptions have been encountered, together with information on velocity and occurrence.

The results are summarized in a theoretical HR diagram with M_{bol} against log $T_{\rm eff}$ (Figure 4-32).

Special emphasis is given to Be stars. Some Be stars are classified as "yes" only because of their strong transient character (mainly in C IV). Comparison of detailed studies of individual stars, however, strongly suggests that we are dealing with the same phenomenon. For example, in spite of the many UV spectra of 59 Cyg, which show high-velocity absorptions with very irregular shapes (e.g., Doazan in Underhill and Doazan, 1982, p. 397), some of the spectra of this star strongly resemble the regular shapes of those in γ Cas as in Figure 4-31 (e.g., Figure 8b in Doazan et al. (1985)). Of the 225 stars without saturated lines, positive evidence for discrete absorption features has been found in 118 stars. We must bear in mind, however, that this number must be regarded as a lower limit because of the transient nature of the lines, especially in the Be stars. For example, at least 50 IUE spectra of γ Cas give no indication of narrow absorptions!

From the distribution in the HR diagram, a number of very significant quantitative results can be concluded:

(a) Among stars brighter than about $M_{\rm bol} = -7$, discrete absorptions are observed at least once in 65 percent of the sample of about 150 stars, counting only those with unsaturated P Cygni profiles, regardless of their effective temperature. There is strong evidence that essentially all supergiants (luminosity class Ia, Iab, and Ib) earlier than B1 show discrete absorption components. In the sample of these bright stars, the occurrence of discrete absorption lines is not correlated with Walborn's "f" or "n" classification, stellar radius, *V sin i*, rotation period, or binarity, nor is it correlated with a reported presence or absence of Hα or X-ray emission for a given star (i.e., the same ratio *yes* to *no* is found among the known Hα or X-ray emitters). However, these two types of

Table 4-6
Well-Known Examples of O, sdO, B, and Be Stars Showing Variable Discrete Absorptions*

Name	HD	Spectral Type	Lines Showing Discrete Components	Typical Velocity (km/s)	References on Variability**
ζ Pup	66811	O4 I(n)f	Si IV	−2200	8,20,21,22,24,29,30
	199579	O6 V((f))	N V	−2300	21
	167771	O7 III(v)((f))	N V, Si IV	−2400	21
15 Mon	47839	O7 V((f))	N V, C IV	−1900	10,21,22
	48099	O7 V	N V, C IV	−2800	21
ξ Per	24912	O7.5 III(n)((f))	Si IV	−2200	19,21,22
δ Cir	135240	O7.5 III((f))	N V, C IV	−2200	21
λ Ori A	36861	O8 III(f))	N V, C IV	−2000	7,21,22
	47129	O8p	N V, C IV	−2200	21
τ CMa	57061	O9 II	N V, Si IV	−2000	21,22
ι Ori	37043	O9 III	N V, C IV	−2000	10,21,22,30
	93521	O9 Vp	N V, Si IV	−150	21
10 Lac	214680	O9 V	N V	−1100	21
α Cam	30614	O9.5 Ia	N V	−1400	21,22
δ Ori A	36486	O9.5 II	N V, Si IV, C IV	−1800	10,21,22,23,30
ζ Oph	149757	O9.5 Ve	N V, C IV	−1400	14,21,22
μ Nor	149038	O9.7 Iab	N V Si IV	−1700	21
ζOri	37742	O9.7 Iab	NV	−1600	21,22
	128220	sdO9	Si IV, CIV	−600	2
ε Ori	37128	B0 Ia	N V, Si IV	−1700	3,21
V861 Sco	152667	B0 Ia((n))	N V, C II, Si IV, Al III, Fe III	−600	13
χ Ori	38771	B0.5 Ia	N V, Si IV	−1400	3,21,22,27
γ Cas	5394	B0.5 IVe	N V, Si IV, C IV	−1400	12
HR 2855	58978	B0.5 IVe	N V, C IV	−200	18
ρ Leo	91316	B1 Iab	N V, Si IV, C IV	−1100	21,22
γ Ara	157246	B1 II	N V, Si IV	−500	21,22
59 Cyg	200120	B1 Ve	N V, CI V	−750	4,6,9,15
ω Ori	37490	B2 IIIe	Si III, Si IV, C IV	−850	16,17,26
λ Eri	33328	B2 III(e)p	C IV	−1000	1
δ Cen	105435	B2 IVe	Si III	−700	25
66 Oph	164284	N2 IV-Ve	Si III, Si, IV, C IV, Al III	−500	1,11,17,18
6 Cep	203467	B2.5 Ve	C IV	−400	1
105 Tau	32991	B3 Ve	C IV	−500	1
HR 7739	192685	B3 Ve	C IV	−500	1
θ CrB	138749	B6 IIIe	Si IV, C IV, Al III	−100	5,7,28

*Typical velocities and representative (but not exhaustive) references regarding variability are given.

**References: 1. Barker and Marlborough (1985), 2. Bruhweiler and Dean (1983), 3. Cassinelli et al. (1983), 4. Doazan et al. (1980), 5. Doazan et al. (1984), 6. Doazan et al. (1985), 7. Doazan et al. (1986), 8. Franco et al. (1983), 9. Grady et al. (1982), 10. Grady et al. (1984), 11. Grady et al. (1986), 12. Henrichs et al. (1983), 13. Howarth (1984), 14. Howarth et al. (1984), 15. Marlborough and Snow (1980), 16. Peters (1982a), 17. Peters (1982b), 18. Peters (1982c), 19. Prinja et al. (1983), 20. Prinja (1984), 21. Prinja and Howarth (1986), 22. Snow (1977), 23. Snow and Hayes (1978), 24. Snow et al. (1980a), 25. Snow et al. (1980b), 26. Sonneborn et al. (1986), 27. Stalio et al (1981), 28. Underhill (1985), 29. Wegner and Snow (1978), 30. York et al. (1977).

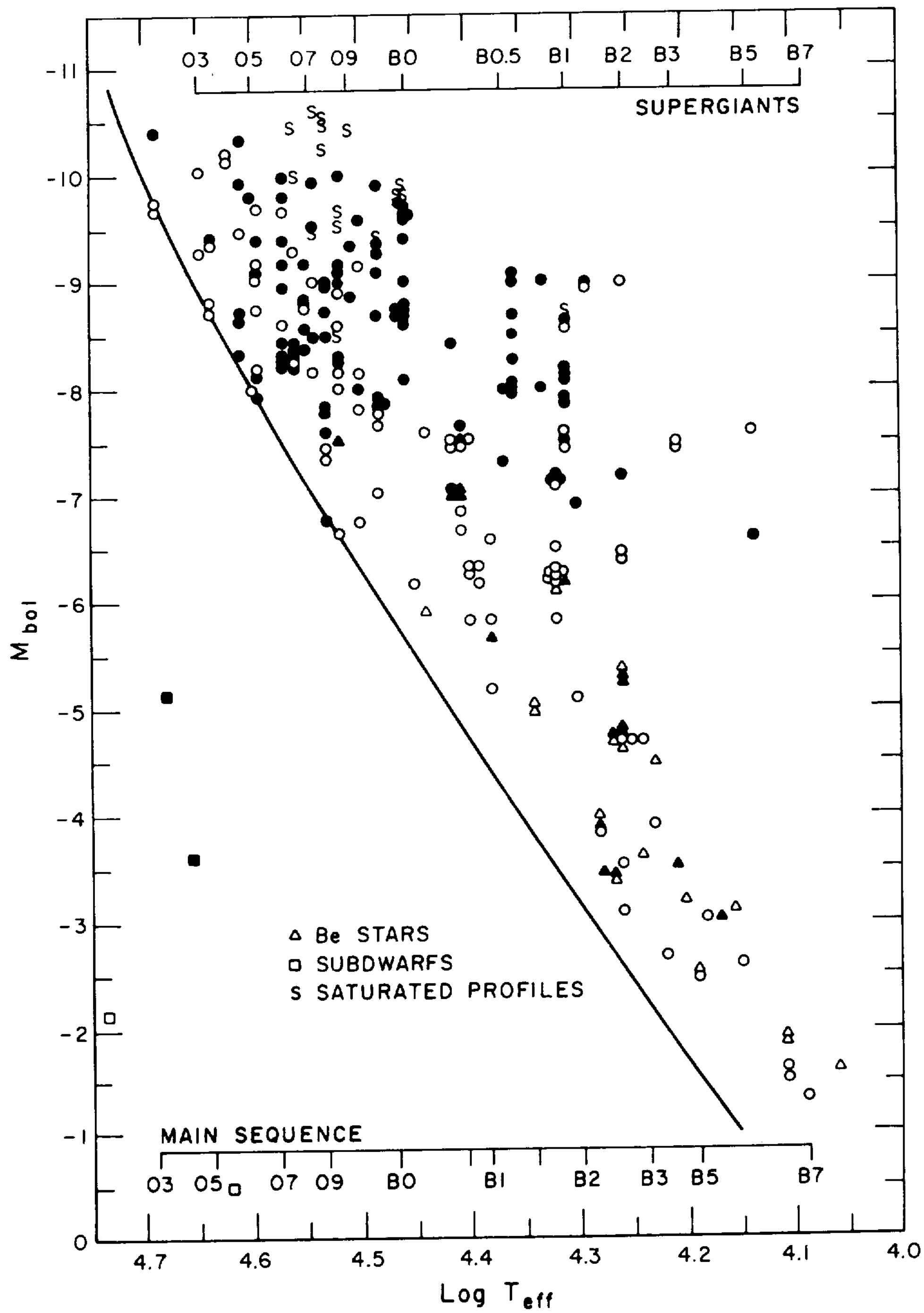


Figure 4-32. A theoretical HR diagram with the position of 241 stars searched for the presence of high-velocity discrete components in C IV, Si IV, and N V (Henrichs and Wakker, 1986): • denotes a positive identification, o means that no spectrum with discrete absorptions has been found with certainty, s denotes saturated P Cygni profiles, making identification impossible, and Δ denotes Be stars. Note that, below $M_{bol} \simeq -7$, Be stars are the only nonsupergiant stars which show discrete absorptions.

emission are known to be varying with time (for example, Ebbets, 1982, and Snow et al., 1981), and almost no simultaneous UV observations exist, which makes this absence of correlation probably insignificant.

(b) For nonsupergiant stars fainter than about M_{bol} = -7, only a fraction of the Be stars shows discrete absorption features. In *none* of the normal B (i.e., non-Be) stars less luminous than M_{bol} = -7 has positive evidence for discrete absorptions been found. This result has been confirmed by Grady et al. (1987b).

(c) In about 50 percent of the Be star sample studied (i.e., 17 of 33), the discrete components are *sometimes* found. This quoted fraction may not be as reliable because, in Be stars, the absorption features tend to disappear completely from time to time. A similar conclusion was reached in a recent study by Grady et al. (1987b).

(d) Of the limited sample of four subdwarfs, two clearly showed evidence of multiple narrow absorption lines. In HD 228220B, the lines are strongly variable. (See Figure 4-36.)

2. Be Stars. The remarkable conclusion that discrete absorptions appear in stars less luminous than $M_{bol} \simeq -7$ (corresponding to spectral type around O9 near the main sequence) *only* if the star is known to be a Be star is probably one of the most pronounced spectral properties of Be stars in the ultraviolet. (Refer to Underhill and Doazan (1982) for a review of general Be star properties.) The C IV doublet appears to be the most clear indicator of this variability (see for example, Barker and Marlborough, 1985). In other words, C IV variability in the UV for stars less luminous than $M_{bol} \simeq -7$ means that the star is a Be star. Note that the reverse is not necessarily true because many Be stars have not yet shown C IV activity. Some Be stars appear to have very complicated UV spectra, making identification of discrete components uncertain. These stars were listed by Slettebak (1982) as showing a shell spectrum in the visible region at the epoch of his observation(s). It would be of interest to see whether the discrete component behavior changes during B–Be–Bshell transitions in the visible.

A search among Be stars for a possible correlation with *V sin i* is presented in Figure 4-33. In this figure, the distribution in *V sin i* of all Be stars with corresponding spectral type (Slettebak, 1982) is also shown. The great similarity between this distribution and that of the sample shows that the statistics are more or less complete. The conclusion is that the fraction of Be stars which show discrete absorption features is probably *not* correlated with *V sin i*. Grady et al. (1987b) essentially confirmed this result on the basis of a larger sample of stars

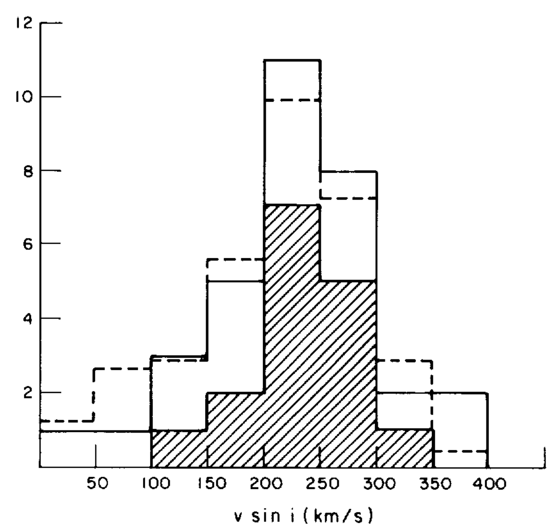


Figure 4-33. The distribution of V sin i of Be stars in the sample searched for high-velocity discrete absorption components (Henrichs and Wakker, 1986): the hatched area denotes positive identification; the dashed line denotes the (scaled) distribution of 136 Be stars (Slettebak, 1982), indicating that the sample is reasonably complete. No correlation with V sin i can be concluded.

and stressed the paucity of discrete components in Be stars with $V \sin i < 150$ km/s. The meaning of such a *threshold* is not established. A closer investigation is worthwhile as to what extent latitudinal effects are important.

3. Discussion. The apparent division between stars brighter and fainter than $M_{bol} \simeq -7$ must tell us something about the physical origin of the phenomenon of discrete UV absorptions, assuming one common origin (which seems likely but is still to be shown). This division apparently does not coincide with the boundary which separates stars with and without detectable mass outflow ($M_{bol} \simeq -6$, Snow and Morton, 1976) because many stars below this line clearly exhibit asymmetric absorption profiles (e.g., τ Sco, μ Col, δ Sco, and X Per) but have never shown indications of discrete absorption components. A systematic study of stars occupying this part of the HR diagram might be worthwhile, especially in view of the well-known uncertainties in M_{bol} (± 0.5 mag is no exception). We must conclude that the presence or absence of a stellar wind as observable from (steady) asymmetric absorption profiles is apparently not related to the occurrence of discrete components.

It is important to realize how the reported variability might influence the foregoing statistics and what its outcome actually implies. For instance, one might wonder if the quoted fraction of 50 percent among the Be stars actually means that all Be stars have discrete components in their spectra during 50 percent of the time. Long-term studies of a few stars (e.g., γ Cas (Henrichs et al., 1983), 59 Cyg (Doazan et al., 1985), and ζ Oph (Howarth et al., 1984)) have taught us that this is *not* the case over a 5-year period. For example, γ Cas was very "active" over a period of 2 years, but during the following 4 years, hardly any discrete absorption feature was found. In 59 Cyg, there are only a few known spectra without the features, whereas in ζ Oph, they vary but never disappear like almost all stars from the study by Prinja and Howarth (1984), which contains only stars more luminous than $M_{bol} = -7$. In summary, it appears that, in Be stars, the features come and go, whereas among the O stars and other B stars, the features are either always present but variable or always absent.

That the occurrence of discrete absorptions depends strongly on the Be character of the stars might, in our opinion, be fundamental for the understanding of the Be phenomenon. Many of the luminous (non-Be) stars which exhibit discrete absorptions also have emission lines in the visual spectrum (at least on some occasions because there are numerous examples of varying emission (e.g., Ebbets, 1982)). It is therefore suggestive to conjecture that the same underlying mechanism that is responsible in some way for the variable optical emission in both Be stars and luminous stars also causes the discrete absorption lines in the UV. It has been suggested that the changing behavior of nonradial pulsations might play a significant role in producing discrete absorptions in the stellar-wind profiles (Henrichs, 1984), analogous to the suggestion by Vogt and Penrod (1983) that nonradial pulsation behavior might be responsible for Be outbursts. (See Section F for possible evidence and further discussion.) Whatever the physical difference between Be and other stars might be, the explanation of the phenomenon of discrete absorption components should, of course, account for the fact that occurrence of the components is not restricted to Be stars.

C. Summary of Properties

This section summarizes the main characteristic features of the discrete absorptions. Significant examples of variabilities in stellar-wind lines in OB stars are given in Figure 4-34 (a-g). Tickmarks on vertical axis indicate zero flux level. The scale is similar to that of Figure 4-32. This summary is based on the work of Lamers et al. (1982), Henrichs (1984), and Prinja and Howarth (1986).

(a) The features are often found in resonance lines of abundant ions (C IV, Si IV, N V, Si III, O VI, and Al III and

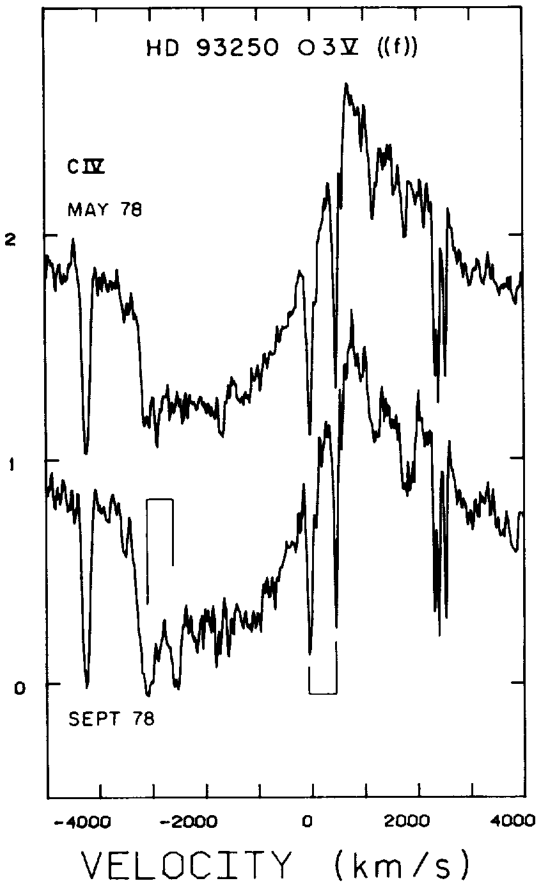


Figure 4-34a. Two spectra in the C IV region of HD 93250, O3 V((f)), taken 4 months apart. Note the sharp additional doublet near -3000 km/s in the lower spectrum.

the nonresonance line of C III). In some stars (e.g., HD 64760, B0.5 Ib), they seem to coexist in Si III, Si IV, C IV, and N V (Figure 4-34). Lamers et al. (1982) claim the simultaneous presence of discrete components over an even wider range of ionization potentials in the star, ϵ Ori, B0 Ia, viz. in Si III, N V, and O VI.

(b) They are sometimes multiple. (Four components or more have been observed.)

(c) The range of the velocity of the center of the feature(s) is usually from a few hundred to 3000 km/s, depending on the star and is typically between 0.5 and 0.9 times the terminal velocity (v_∞) of the stellar wind, if the latter can be observed (as a steep edge of a saturated P Cygni profile). There is no correlation between this fraction and the actual value of the terminal velocity. In general, the later the spectral type, the lower the velocities involved. In the star, θ Crb, B6 IIIe, the central velocity can be as low as around zero (Doazan et al., 1984).

(d) When the features are observed at high velocity and are found in lines of different ions, they occur at the same velocity within the observational error

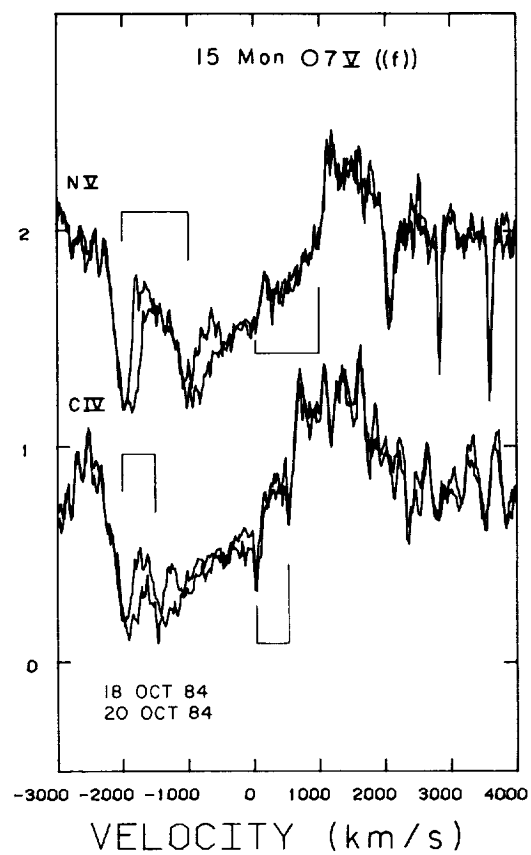


Figure 4-34b. Significant variability within 2 days in N V and C IV lines in 15 Mon, O7 V((f)).

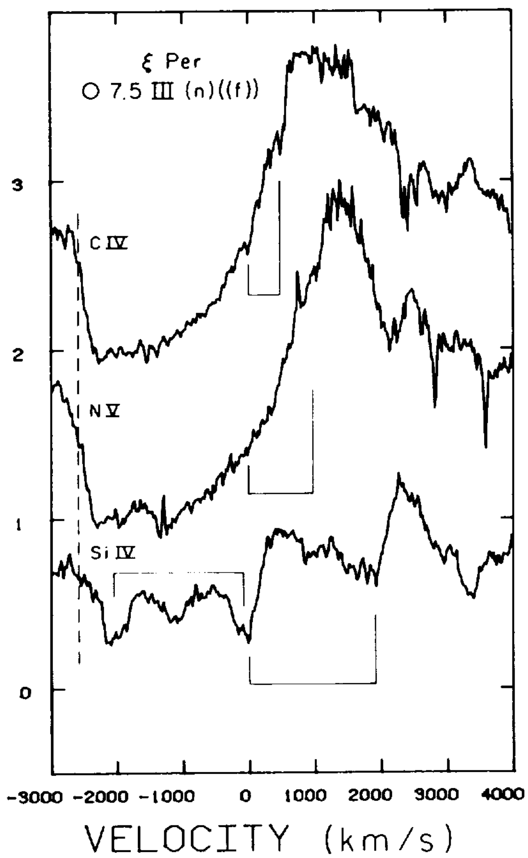


Figure 4-34c. Resonance lines in ξ Per, O7.5 III(n)((f)): dashed line denotes the terminal velocity derived from the C IV and N V profiles. The discrete components in the Si IV doublet are found at significantly lower velocity. See Figure 4-44 for the time series of this star.

(about 20 km/s). In particular, the velocity does not correlate with ionization potential.

(e) The width is typically 30 to 300 km/s, or in other words, between 0.03 and 0.14 v_∞ (Prinja and Howarth, 1986), again without depending on the value of v_∞ itself. Although differences in width between lines of different ions have been found, there is no substantial evidence that a systematic difference exists. The width is not correlated with the central depth of the feature.

(f) A typical observed value for the column density (not corrected for abundance or ionization fraction) is 2×10^{14} cm^{-2} with a range of a factor of 5 on both sides (Figure 4-35). Interestingly enough, the column densities of the discrete components do not depend on those of the underlying P Cygni profile (Figure 4-35). In other words, the column density of a discrete component in an early O star is often comparable to that of a B or Be star. One must bear in mind that a specific model has been used to derive the quoted column densities. However, different methods have thus far given similar results, and the

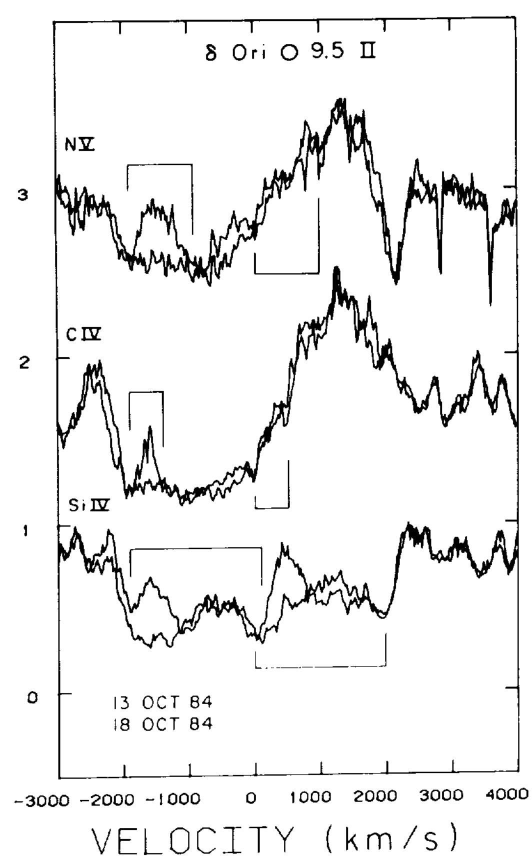


Figure 4-34d. Two spectra of δ Ori, O9.5 II, taken 5 days apart, which is approximately the binary period, illustrating that the variability is not binary-phase-dependent.

range quoted is therefore considered to be fairly representative.

(g) The degree of ionization in the discrete components is somewhat complicated. Lamers et al. (1982) found that the degree of ionization is higher than that of ordinary wind lines. This conclusion has been supported by the study of Prinja and Howarth (1986). However, these authors realized that, if one considers the effects of Auger ionization, the simplest explanation of the data is, in fact, that the level of ionization in the discrete components, as measured by the dominant ions, is lower than in the ambient wind. The efficiency, however, with which superions (N^{4+}, O^{5+}) are produced in the discrete components is enhanced. (More study is obviously required, particularly in the form of simultaneous X-ray and ultraviolet observations.) Regarding the time-dependence of the ionization conditions in discrete components, Prinja and Howarth found that, for most of the stars in their sample, the column density *ratio* of the components of different ions did not change significantly as a function of time. This led them to conclude that the variability observed in the strength of the discrete absorptions is more likely attributable to fluctuations in the total (hydrogen) column density than to changes in the local ionization balance.

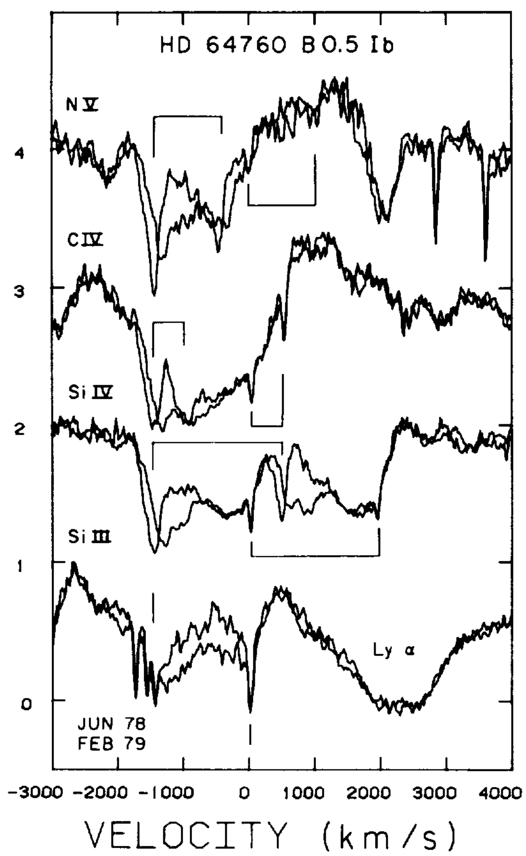


Figure 4-34e. Two superposed spectra of HD 64760, B0.5 Ib, ions arranged according to ionization potential, illustrating the simultaneous presence (and variability) of discrete components over a wide range in ionization. The second spectrum was taken 8 months after the first.

(h) Some stars display a correlation between velocity and column density: the higher the velocity, the weaker the discrete component. This correlation was called a "memory" by Henrichs (1984) in order to acknowledge the fact that, for a given star at a given episode, the discrete absorptions might completely disappear but that, when they come back, the relation between the new velocity and column density tends to be the same as before. The range is about 10 percent in velocity and a factor of 5 to 10 in column density. It should be emphasized that there are many more stars without such a memory. Prinja and Howarth (1986) found that, in HD 199579, O6 V((f)), the correlation is also present, but instead, a lower column density is related to a *lower* velocity.

(i) The strength and velocity are often variable on time scales of hours, days, and weeks. However, not many detailed observations with high time resolution

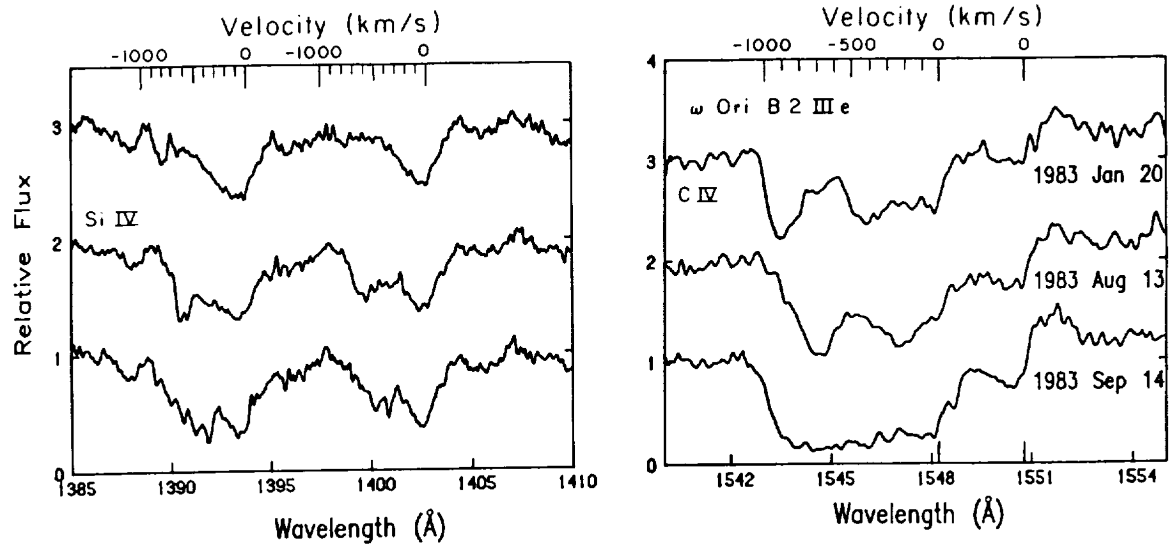


Figure 4-34f. Different types of profiles in ω Ori, B2 IIIe (Sonneborn et al., 1987). The behavior in Si IV is similar but not identical to that of C IV.

(i.e., hours) exist. No periodicity in the variability has ever been found. In the detailed study of 21 stars by Prinja and Howarth (1986), it was found that, for a given star, the column density can vary by at least a factor of 2 and at most by a factor of 10, with an average of 3. The typical time scales are days and are comparable to the dynamical time scales of variations in Hα emission as discussed by Ebbets (1982). In γ Cas, B0.5 IVe, time-resolved observations covering a couple of days have shown that the velocity of a *high*-velocity feature does not change but that its strength decreases. In ξ Per, O7.5 III(n)((f)), however, the features are found to accelerate, "settle" at high velocity, and consequently become fainter. (See Figure 4-44b and Section D for more details.) Two time scales are needed to describe the variability: (1) a short time scale (hours) identified with the flow time scale, and (2) a longer time scale (weeks, months) to characterize the strength of each episode, varying from completely absent to permanently present.

(j) In a few well-studied cases (O stars ξ Per and δ Ori and most of the Be stars), *low*-velocity absorption is sometimes associated with the high-velocity component(s). (See for example, Figures 4-34 and 4-44.) The best-studied star thus far, ξ Per, showed that the low-velocity features evolved into high-velocity absorption within 1 day (Figure 4-44). It remains to be shown, however, whether the behavior of ξ Per is representative or not. The low-velocity features can be asymmetric or have irregular shape and are clearly much more variable than the high-velocity features. There appears to be a general trend that the later the spectral type, the more pronounced the low-velocity features. When observed simultaneously in different ions, the shapes may differ for the individual ions, which therefore results in different average velocities. In contrast to the case of ξ Per, in θ CrB, B6 IIIe, the star with the latest spectral type in which the variable absorption features are reported, the *averaged* central velocity is found to be slowly increasing from around 0 to −140 km/s over a period of

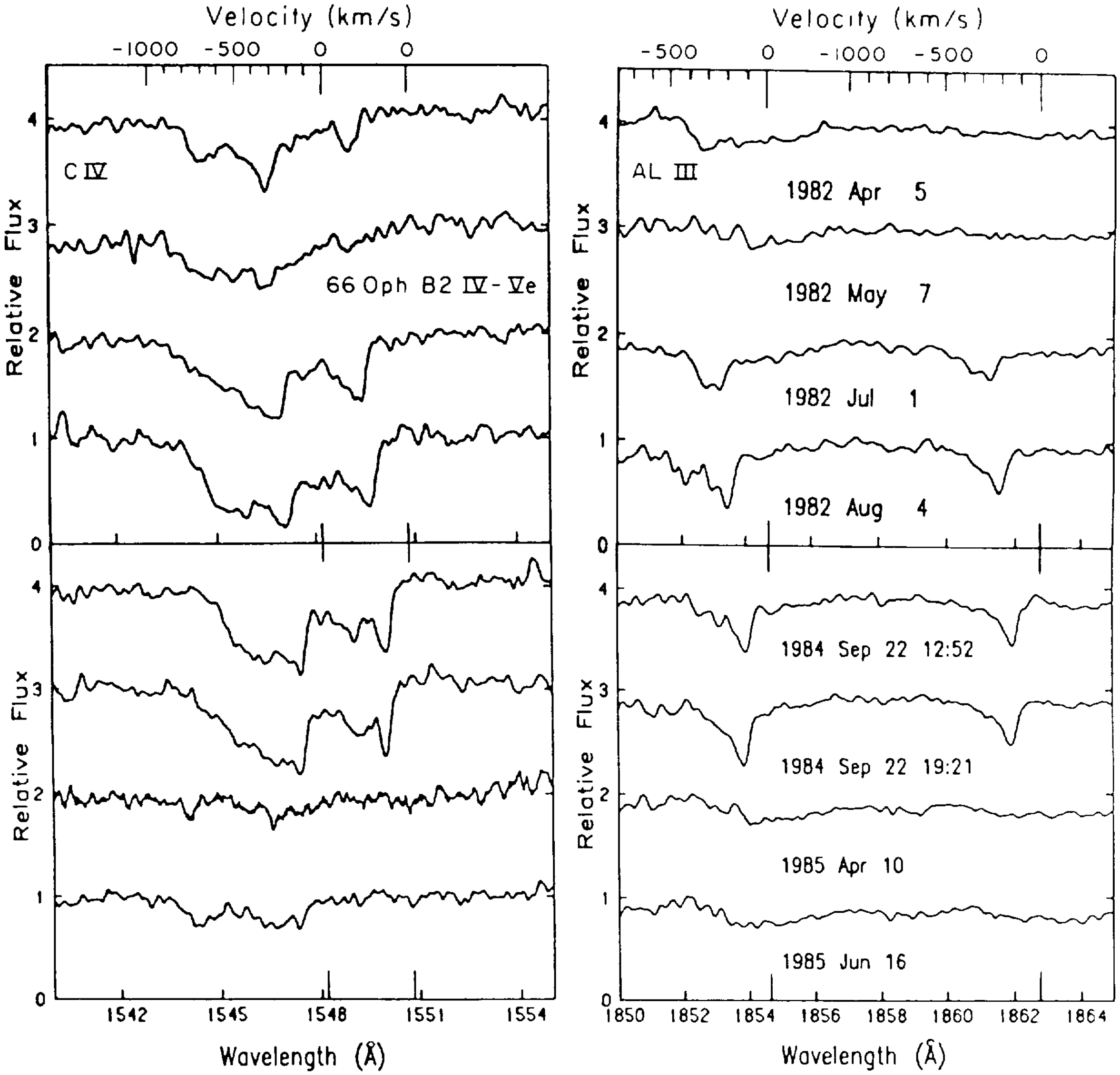


Figure 4-34g. Variability in the resonance lines of C IV and Al III in 66 Oph, B2 IV–Ve (Grady et al., 1987a). Note that the character of the profiles is very different in the two ions and that the absorptions disappeared in 1985.

5 years (Doazan et al., 1984). However, the behavior on a daily time scale (or shorter) of the absorptions in this star is not known. Again, whether such behavior is typical for a Be star or not is an open question. Obviously, such observations are extremely important because they might reveal how the features are actually formed and/or whether the characteristics are epoch-dependent.

D. Variability

Numerous examples exist of varying high-velocity absorption in UV line profiles of many stars on time scales down to 30 minutes. Typical examples are listed in Table 4-6. The list is not

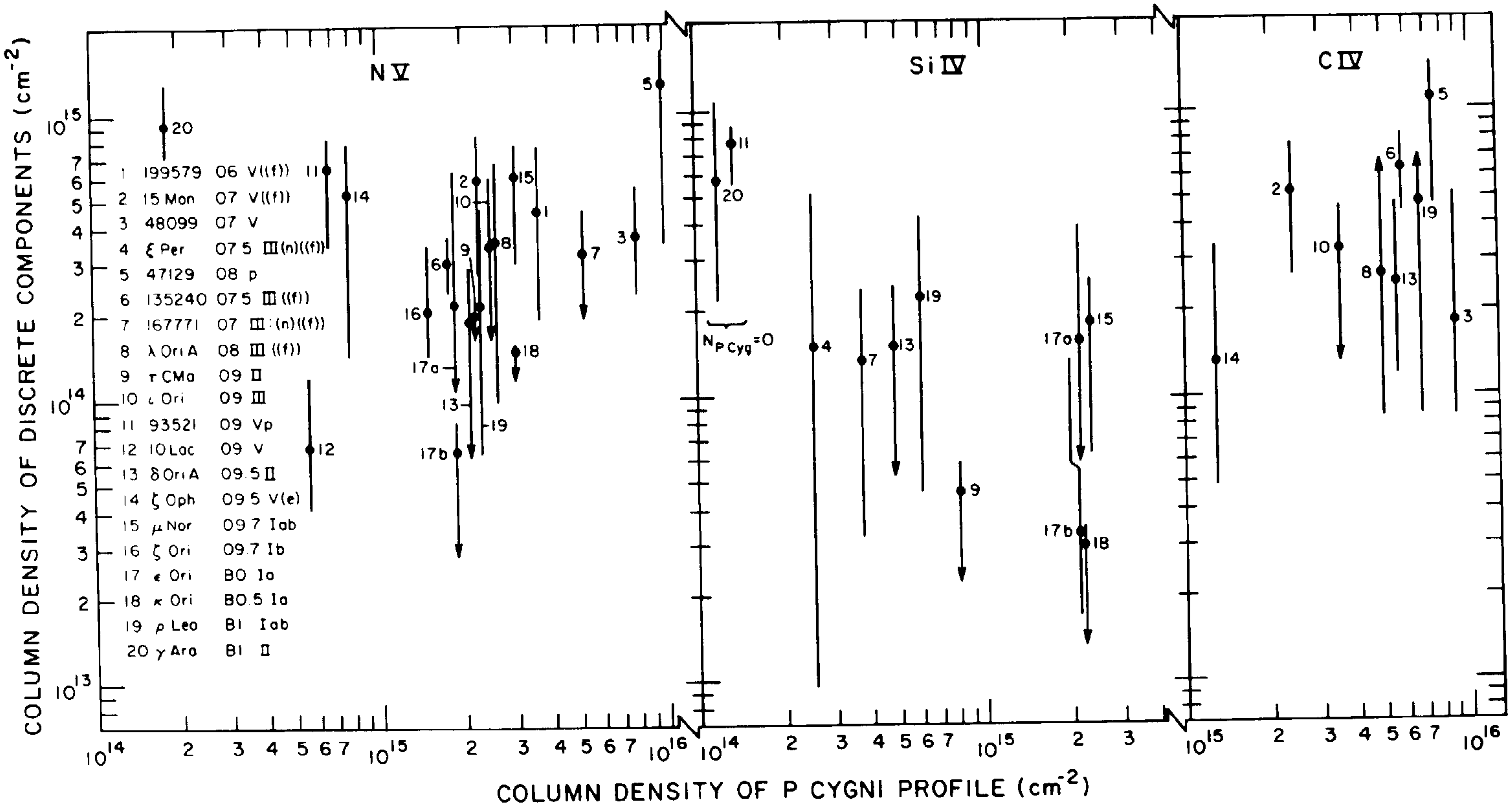


Figure 4-35. Observed range in column density of discrete components in C IV, Si IV, and N V as a function of column density of the underlying P Cygni profile (based on Prinja and Howarth, 1986): dots denote average values, bars indicate highest and lowest measured value (no error bars!), and arrows up and down indicate that saturated discrete components or no components at all had been detected, respectively. Star 17 (ε Ori) has two distinct sets of discrete components, indicated by a and b, respectively. From this graph, one can conclude that the column density of discrete components is not related to the column density of the P Cygni profile or the spectral type.

exhaustive, but rather illustrative. Earlier examples can be found in Snow (1979), Underhill and Doazan (1982), Marlborough (1982), and Henrichs (1984). It is important to notice that not only the high-velocity absorption part is variable, but that changes are also observed at lower velocities. Most of the papers report only qualitatively on variable absorption. Some of the papers cited in Table 4-6 contain quantitative measurements of column density and velocity of especially the discrete components, based on many of high-resolution IUE spectra of these stars, notably the paper by Prinja and Howarth (1986). The main conclusion is that variability in strength and velocity is most likely a property of all discrete absorptions.

1. Quantitative Measurements. Results of quantitative measurements of properties of discrete absorptions depend on the model and method used. Therefore, we summarize briefly the different techniques and their underlying assumptions before quoting results. The technique for deriving the column density of narrow absorption lines, including the case of multiple components, is described in detail in Henrichs et al. (1983). The basic assumptions are that the absorption line is formed in a homogeneous plane parallel slab of gas between the star and observer, that forward scattering and emission are negligible, and that no radiative coupling occurs between the line and the (often present) underlying P Cygni profile. (The latter assumption was not made by the analysis of Lamers et al. (1982), but we note that the resulting column densities are similar.) Such an analysis is basically similar to interstellar work. The farther the gas layer is from the star, the more reasonable the assumptions are likely to become appropriate. These assumptions imply that the intensity of the radiation emerging from the star, including stellar wind (I_o), is modified according to:

$$I(\mathrm{v}) = I_o \exp\left\{ -\tau_c \exp\left[-\left(\frac{\mathrm{v} - \mathrm{v}_c}{\mathrm{v}_t} \right)^2 \right] \right\} , \qquad (4\text{-}39)$$

where v is the velocity with respect to the stellar rest frame, τ_c is the optical depth at the center of the line which has a radial velocity v_c, and v_t is the broadening parameter (v_t = 0.601 FWHM) characterizing the Gaussian velocity dispersion. A least-square fit of such a profile (corrected for the instrumental profile of the spectrograph used) returns the values of τ_c, v_c, and v_t. Such a prescription gives for the column density:

$$N = \frac{\sqrt{\pi}}{\pi e^2/mc} \, \frac{1}{\lambda_o f} \, \frac{\tau_c \mathrm{v}_t}{1 + \mathrm{v}_c/c} , \qquad (4\text{-}40)$$

where $\pi e^2/mc$ = 0.02654 cm^2 s^{-1}, λ_o is the laboratory wavelength of the line which has oscillator strength f, and c is the velocity of light. In the case of doublets, it is assumed that the value for the broadening parameter v_t is the same for both members. This implies that

$$\frac{\tau_1}{\tau_2} = \frac{\lambda_1 f_1}{\lambda_2 f_2} \qquad (4\text{-}41)$$

for both optical thick and thin lines. The use of this well-known doublet ratio greatly improves the reliability of a fit. (Note that the question of optical thickness arises only when equivalent width measurements are to be converted into column densities, which is avoided by the present method.) When a doublet is being fit and the velocity scale used is relative to one of the doublet lines, the velocity of the second doublet line must be modified accordingly. (See Henrichs et al., 1983, Equation (3.3).) When multiple (overlapping) discrete components are present, profiles of the first type can be combined where the same set of assumptions apply.

In practice, a basic difficulty of this method is to separate contributions from the "persistent" part of the P Cygni profile, including the photospheric lines, and the superimposed discrete absorption(s). In the case of γ Cas (Henrichs et al., 1983) and ζ Pup (Prinja, 1984), spectra without components (e.g., Figure

4- 31) of the same star were used as a reference spectrum relative to which the measurements were done. In most stars, however, no such spectra are available. Therefore, in V 861 Sco (Howarth, 1984), ζ Oph (Howarth et al., 1984), and ω Ori (Wakker and Henrichs, 1987), a "theoretical" P Cygni profile (e.g., as described by Castor and Lamers, 1979) was fitted to the part of the resonance lines outside the discrete absorption region, and the column densities were derived with respect to this local background. The complicated structure of the Si IV doublet in ξ Per (see Figure 4-44), classified as O7.5 III(n)((f)) by Walborn (1972) and as O7.5 I by Conti and Leep (1974), was analyzed by Prinja et al. (1984) by using a spectrum of λ Ori. This star has a spectral type comparable to that of ξ Per, but it often shows only photospheric Si IV absorption lines and rotates much more slowly. After the convolution of such a λ Ori spectrum with a proper rotational broadening profile, the constructed spectrum (containing many photospheric lines mainly due to Fe IV and V) appeared to be very similar to that of ξ Per, except near the Si IV doublet, and was used as a photospheric input spectrum in constructing the template used to separate the persisting and varying part in ξ Per. In a similar way, Prinja and Howarth (1986) used the C IV part of the spectrum of the O7 V((f)) star, 15 Mon ($V \sin i$ = 63 km/s), as a photospheric standard for the analysis of the O7 V star, HD 48099. They also used the Si IV region of λ Ori, O8 III((f)); μ Col, O9.5 V ($V \sin i$ = 97 km/s); and τ Sco, B0.2 V ($V \sin i \leq 10$ km/s) for analyzing other stars. Figure 4-35 gives an impression of values of column densities thus obtained.

It is appropriate to stress that many ad hoc assumptions were made in the foregoing analyses, most of which were made only because one did not know how to do better. The justification of this kind of analysis is found in the many cases in which the obtained fits represent the observed profiles remarkably well. In other cases, especially in the later Be stars, the irregularity of the observed profiles clearly indicates the inadequacy of the described method.

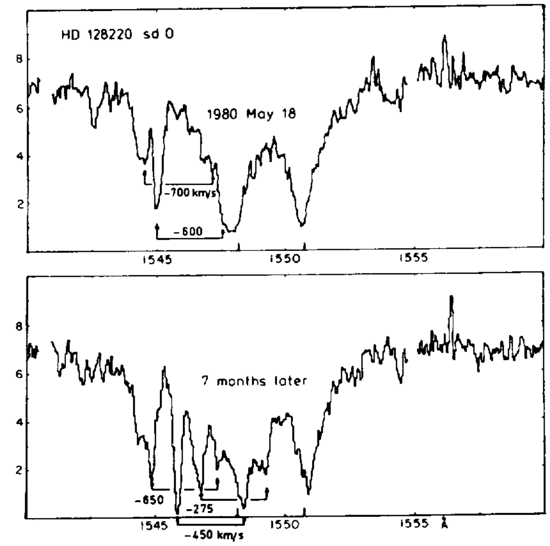


Figure 4-36. Two IUE spectra in the C IV range of the O subdwarf, HD 128220B (after Bruhweiler and Dean, 1983; graph courtesy of Giddings et al., 1986). Note the striking variability in strength, velocity, and multiplicity.

We present here some remarkable properties of the narrow lines which emerged from the few existing detailed studies of individual stars. We review, in turn, the "memory" found in some stars and the time scales involved.

2. The Memory. Figure 4-37 displays the *memory* of γ Cas, B0.5 IVe, regarding the high-velocity narrow absorptions found by Henrichs et al. (1983): a higher velocity is clearly correlated with a smaller column density in spite of the fact that, between the different episodes, the narrow absorptions sometimes disappeared completely. This is valid for all three ions in which they have been observed. However, Doazan (private communication) reports that spectra of γ Cas in 1986 indicate that the described memory does not hold when the new epoch is included, possibly implying that the phenomenon might be time-dependent.

A good example of a memory was found in ω Ori, B2 IIIe, by Sonneborn et al. (1987) and Wakker and Henrichs (1987). Figure 4-38 is an example of the Si IV doublet in ω Ori. The four

C-4

spectra shown in this figure are arbitrarily chosen and have no particular time sequence. The range in velocity is again rather small but very significant and occurs simultaneously in the Si IV, C IV, and Si III (λ1206) lines. Figure 4-39 shows this memory of the discrete absorption component of the Si IV and C IV lines in a quantitative way. Note that these narrow lines are present in all spectra of ω Ori (also in the Si III λ1206 line) and that sometimes the (broader) low-velocity absorption part of the profile is also variable.

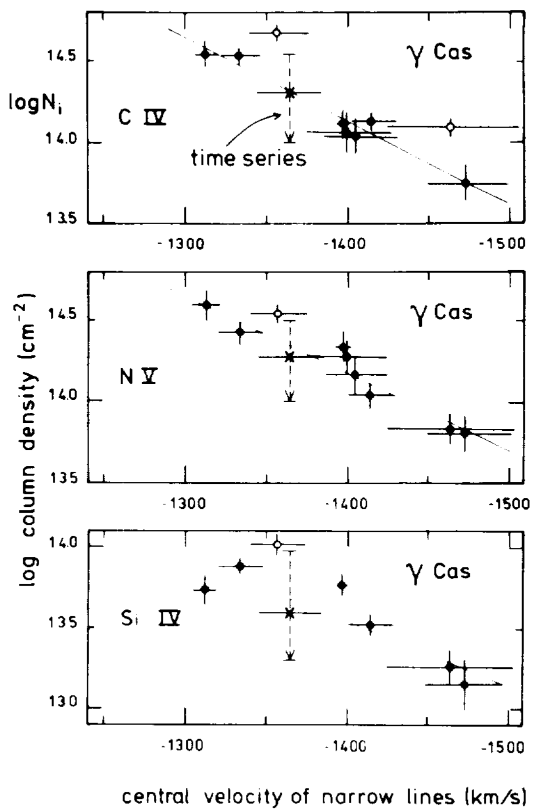


Figure 4-37. The memory in γ Cas (Henrichs et al., 1983). A higher column density is clearly correlated with a lower velocity in spite of the fact that, between different episodes, the discrete components might completely disappear. Open circles denote uncertain points due to instrumental effects. The vertical dashed lines are the ranges observed in six spectra taken within 15 days (see also Figure 4-43).

A third example of a similar memory is found in ξ Per, O7.5 III(n)((f)), by Prinja et al. (1984, 1987). (See Figure 4-40.) Note that, in this case, the velocity (in the stellar rest frame) is in the range of 2000 km/s, much higher than in the previous two stars.

A different kind of memory, found in ζ Pup, O4 I(n)f, by Prinja (1984), is illustrated in Figure 4-41. In this case, there are often two absorption components at different velocities, but when the first is strong, the second is weak and vice versa. Two more examples, 10 Lac, O9 V, and, HD 199579, O6 V((f)), are shown in Figure 4-42 (after Prinja and Howarth, 1986). Remarkable is the thus far unique case of HD 199579, which also seems to have a significant correlation, but with a positive slope (i.e., a higher (negative) velocity is related to larger column density), as opposed to the "memories" of the stars previously mentioned.

It is important to mention that there are many more well-studied stars for which no memory of any kind has been established (e.g., ζ Oph, O9.5 V(e) (Howarth et al., 1984), and 66 Oph, B2 IV–Ve (Barker and Marlborough, 1985), and most of the stars from Prinja and Howarth's study. There exists a very long record of the B1 Ve star, 59 Cyg (e.g., Doazan et al., 1985, and references therein). Because of the complexity of the structure of the absorption features in this star, only equivalent width measurements have been made, so that no conclusion about a memory can be made.

We should mention that there is no clear explanation for the existence of the described behavior. An obvious difficulty is that, in some stars, the significance of the phenomenon is very high, whereas in other stars, no trace of such behavior can be found. More discussion of this remarkable phenomenon will appear in Section E after the available detailed time-sequence studies of these stars have been described.

3. Time Scales. In IUE spectra of γ Cas taken in May 1978, Henrichs et al. (1980) noticed that discrete absorptions were clearly present in a spectrum taken 7 days after a spectrum without

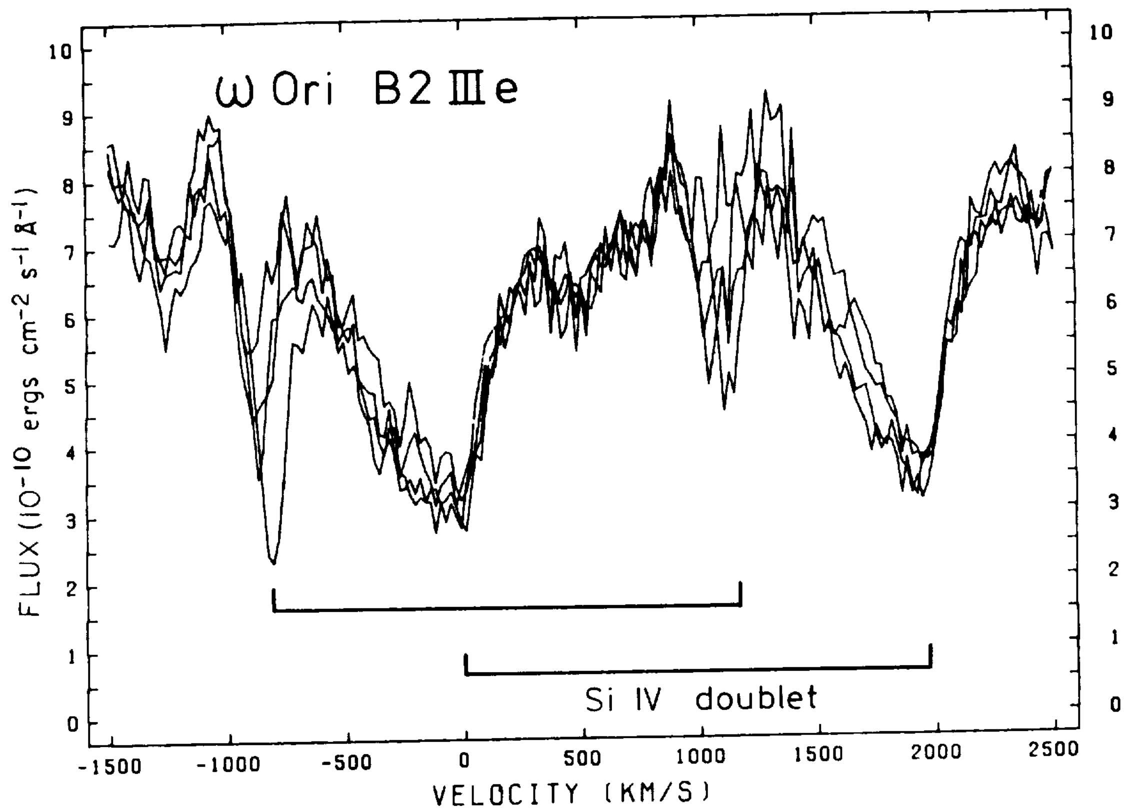


Figure 4-38. The memory in the Si IV doublet in ω Ori (Wakker and Henrichs, 1987). Although there is no particular time sequence in the four spectra, a higher velocity is correlated with a weaker line.

any trace of them. This observation led these authors to suggest that the appearance of discrete components might be caused by ejected material that has a higher density than the ordinary stellar wind. This material would be accelerated (by radiation pressure) during a short time (typically less than 1 day) up to its (own) terminal velocity, causing a narrow absorption line at high velocity. This short acceleration time would make it rather unlikely to find the lines at low velocities, which would explain the frequent absence at low velocity. Subsequent expansion of this dense layer (with a constant velocity) would then cause the column density of the narrow lines to decrease as a function of time proportional to $(t_0 + t)^{-2}$ where t_0 is a constant. (See Section E.3 for a detailed description of this model.) With these time scales in mind, new observations of γ Cas were made in the hope of recording the appearance and acceleration of the narrow absorptions and its decay. Figure 4-43 (from Henrichs et al., 1983) displays the result: a clearly decaying narrow absorption line on a time scale of 1 week with hardly any change in velocity, in accordance with the expectation. The range in column density is indicated in Figure 4-37. The appearance of the discrete absorptions, however, has never been found in this star, except that in one spectrum of γ Cas (March 25, 1980), there is a very broad and strong absorption component at low velocity (–650 km/s as opposed to the narrow lines at –1400 km/s a few days later). A 48-hour continuous observation run in January 1982 of both γ Cas and 59 Cyg (Grady et al., 1982) showed, unfortunately, only very weak narrow absorptions at high velocity in these two stars. Since then, γ Cas seemed to have been inactive

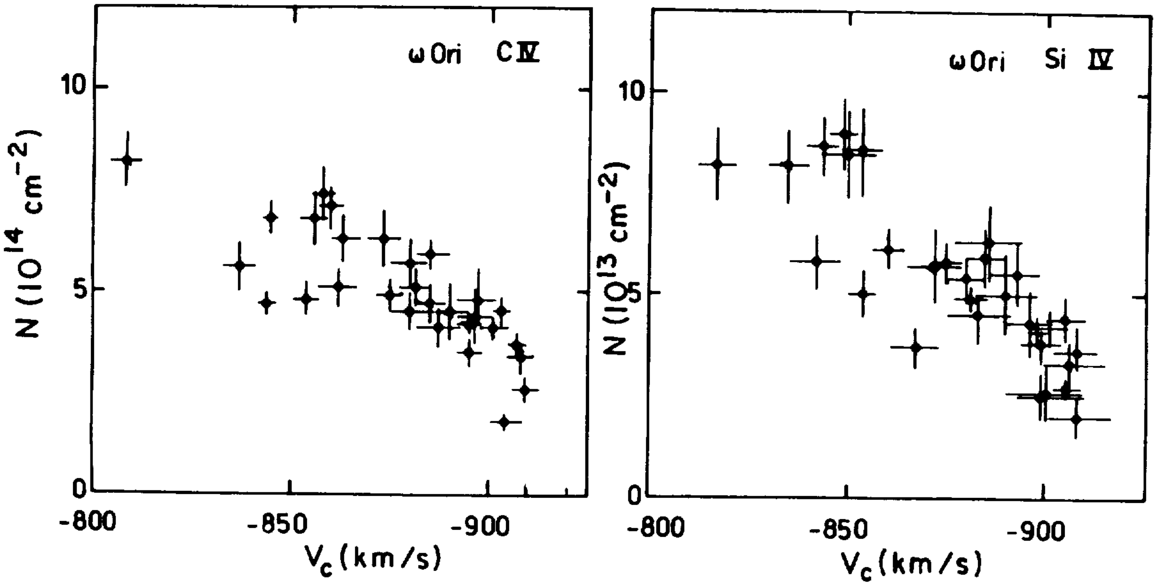


Figure 4-39. The memory in the C IV and Si IV lines in 28 spectra of ω Ori (Wakker and Henrichs, 1987) taken over a 3-month interval. Formal error bars, obtained from least-squares fits, are indicated. A graph for the Si III λ1206 line looks very similar.

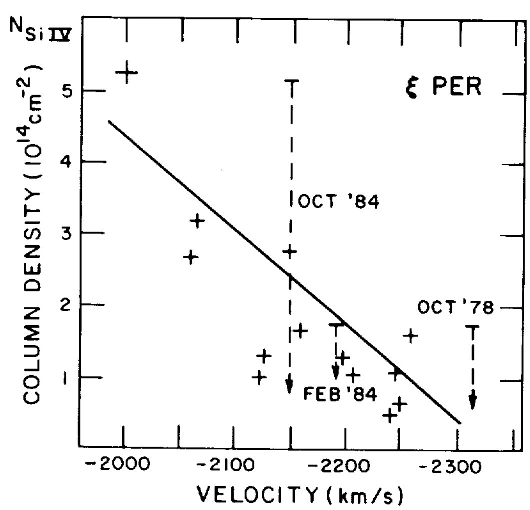


Figure 4-40. The memory of the discrete high-velocity absorptions in ξ Per (Prinja and Howarth, 1986). The ranges in column density of the three time series (Figures 4-44, 4-45, and 4-46) are indicated. The dashed line of the October 1984 data above the solid line (eyeball fit) represents 6 hours of data; below the line it pictures the next 24 hours.

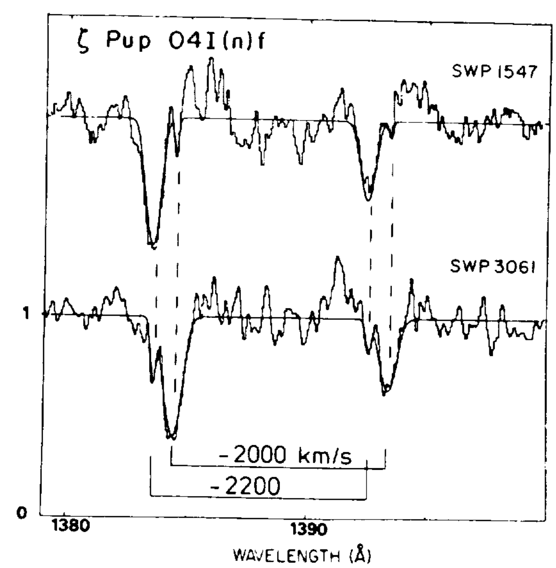


Figure 4-41. A different memory in ζ Pup (Prinja, 1984). Often two separated components are present, but if the high-velocity component is strong, the low-velocity component is weak and vice versa. Here the persistent part of the SI IV line is removed.

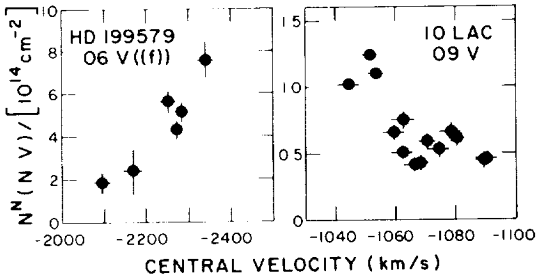


Figure 4-42. The memories (?) of HD 199759 and 10 Lac (Prinja and Howarth, 1986). In the first star (the only one known), the correlation has an opposite sign. Such an apparent correlation appears to be present only rarely.

and there were, with two exceptions, no spectra with narrow absorptions until the summer of 1986, when strong absorption around -1400 km/s reoccurred (Doazan, 1986). This star is a very good example of the two time scales involved: a long time scale (months) characterizing the presence of the discrete components and the short time scale characterizing the variability of the features when they are present.

Much more fortunate were the observations of ξ Per, O7.5 III(n)((f)), reported by Prinja et al. (1984). An extended paper was given by Prinja et al. (1987). In September 1983, a broad low-velocity feature seemed to have almost disappeared in 7 hours, while at the same time, the narrow feature at high velocity had become considerably stronger. (See Figure 6 in Henrichs (1984).) A second successful attempt to pin down the short time scale was made in February 1984. (See Figure 4-44.) Four spectra taken within 2 hours showed a gradual increase in strength of an appearing and accelerating broad low-velocity feature, whereas the high-velocity narrow absorption continued to decrease in strength. An exact pattern was found in spectra of ξ Per taken in October 1978 (available through the IUE data bank). Encouraged by these results, Prinja et al. (1987) monitored the star over several days (Figure 4-44) with varying time resolution down to 0.5 hour. These observations showed, for the first time, a complete cycle of the formation and development of a high-velocity narrow component in ξ Per, beginning with a broad low-velocity feature. The general picture emerging from these observations is that, first, a rather broad absorption feature centered around -1300 km s^{-1} begins to develop (superposed on the P Cygni doublet). Figure 4-44b shows the velocity measurements. At the beginning of such an episode, the red emission of the P Cygni profile is somewhat enhanced. The broad absorption feature first grows in strength and, consequently, strongly narrows during its acceleration, which takes about 0.5 day. After becoming a narrow feature, it remains almost stationary in velocity at about -2150 km s^{-1}. From this point, the strength (column density) of the discrete absorption line decays rapidly with time proportional to $(t_0 + t)^{-2}$, where t_0 is a constant (Figure 4-45). One can see that another episode begins to develop around October 20 at 8:37 UT. On two previous occasions, similar behavior was observed (Figure 4-46), but with

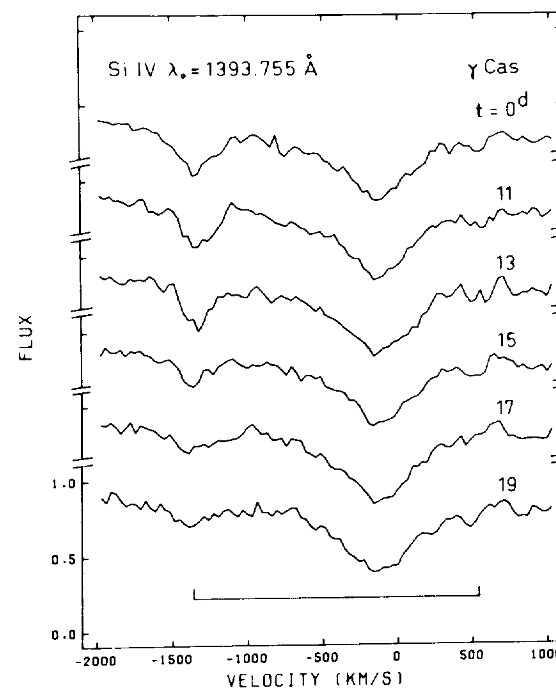


Figure 4-43. A time sequence of γ Cas spectra in 1980 (Henrichs et al., 1983). A time axis is added. Notice the strong decay of the narrow absorptions without significant change in velocity.

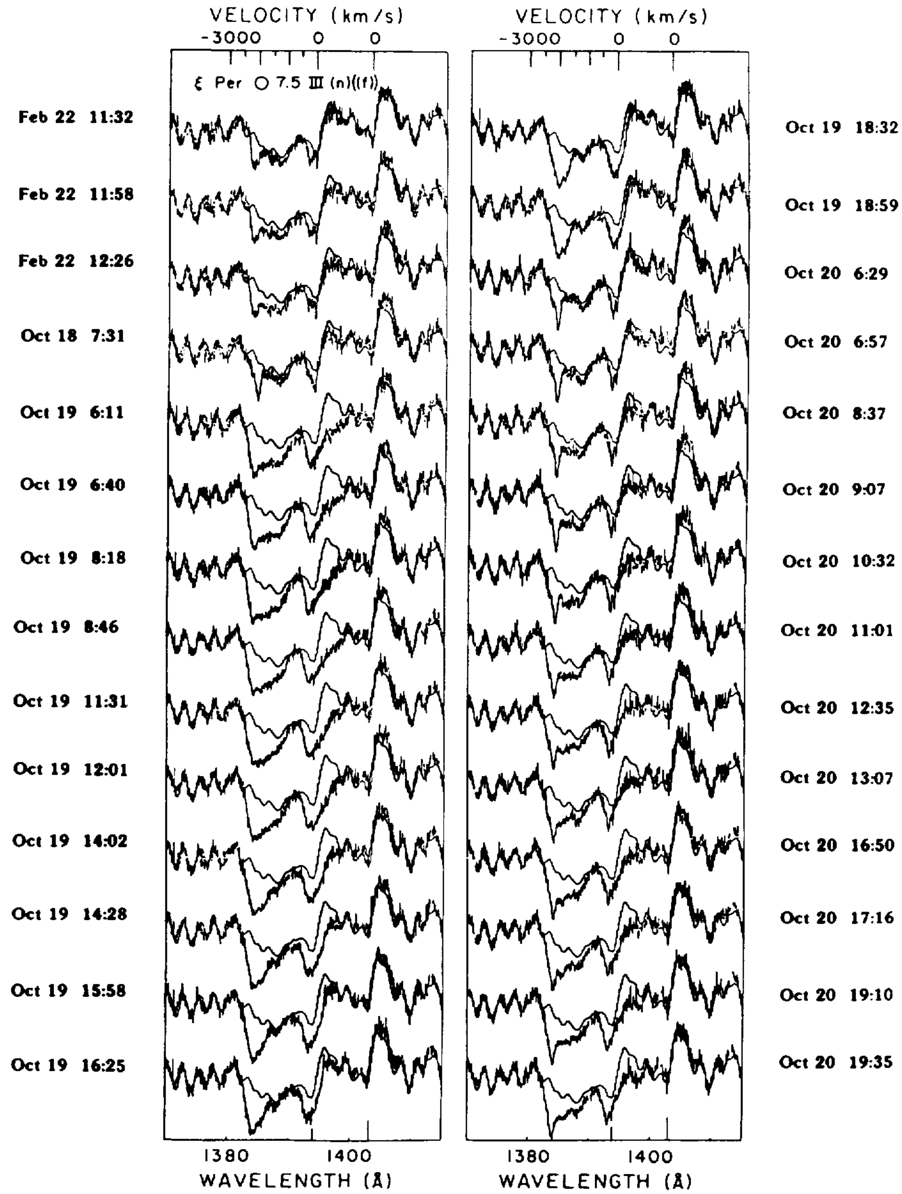


Figure 4-44a. A complete cycle of the formation, development, and decay of discrete components in ξ Per, O7.5 III(n)((f)) (Prinja et al., 1987). The first time sequence was obtained within 1 hour in February 1984 (Prinja et al., 1984); the second within 2.5 days in October 1984 with a time resolution of 0.5 hour over two blocks of 12 hours. The rest wavelengths of the Si IV doublet are marked. A template spectrum used to separate varying and persistent parts of the spectrum (see text) is superposed on each spectrum as a reference. Note the gradual increase of additional broad absorption near λ1388 in the first three spectra. This also happened between October 18 and 19—a very broad low-velocity feature (present in both doublet members) appeared subsequently narrowed when it accelerated during the next 12 hours (October 19 from 6:11 to 18:59). The final velocity was about -2150 km/s, significantly less than the edge velocity of the completely saturated C IV and N V profiles. Note that, in the first October 20 spectrum, the emission *is strongest and is decreasing thereafter. This emission phase signals the beginning of a new cycle, visible as new broad absorption around 1388 Å. The column densities are displayed in Figure 4-45.*

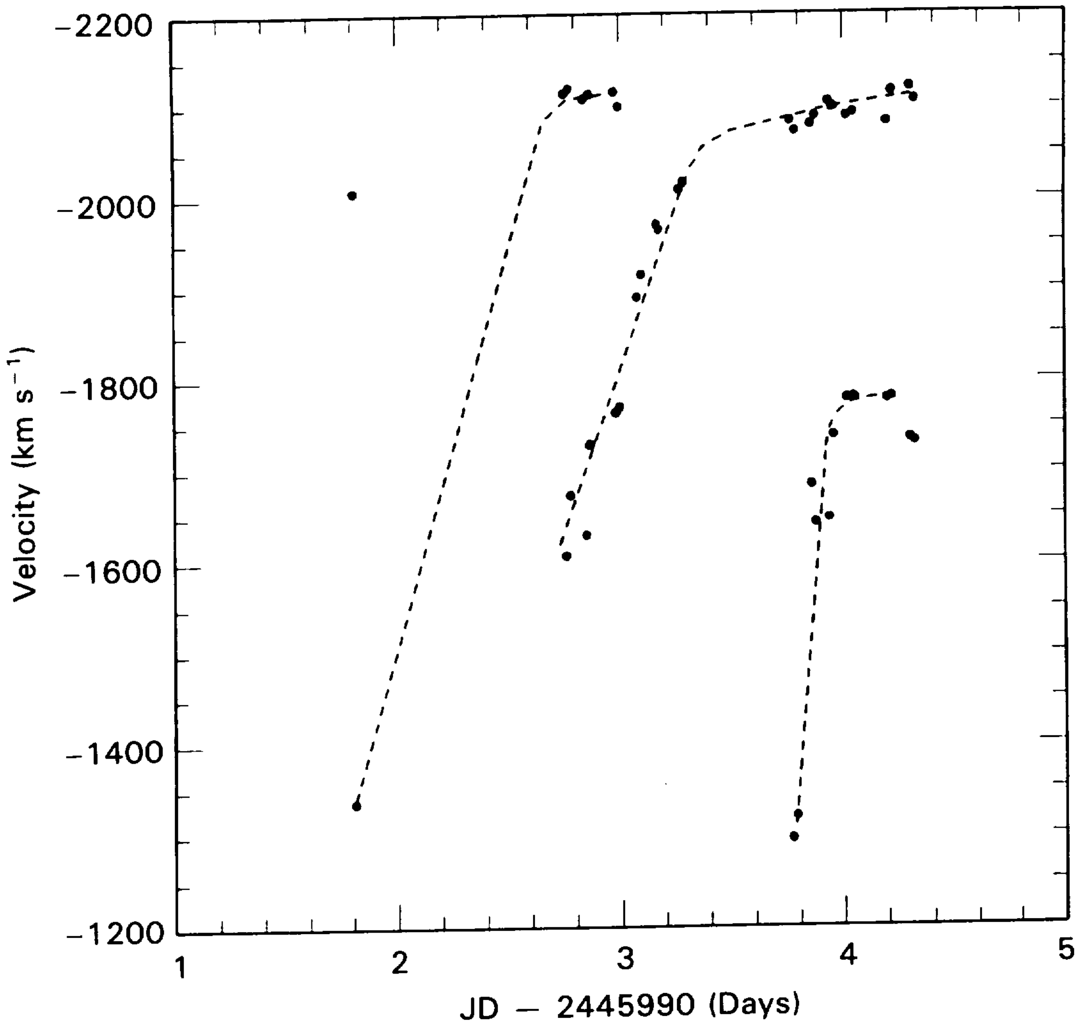


Figure 4-44b. Central radial velocity of the discrete absorption components in ξ Per as a function of time corresponding to Figure 4-44a (Prinja et al., 1987). The velocities were measured on spectra which are rectified using the procedure described in the text. The dashed lines are suggested connections between the different points. There are not enough data available to conclude a possible periodicity. The acceleration of the feature is pronounced.

different central velocities, different values of t_0, and different proportionality factors. It is interesting to note that the latter contains information on how much matter is involved with such an event. A preliminary conclusion is that the more matter is involved, the lower the central velocity of the discrete absorption line is at its maximum velocity. Assuming a spherical shell (for simplicity), Prinja et al. (1987) derive a typical mass of $6 \times 10^{-8}\ M_\odot$ for the October 1984 event. Before firm conclusions can be drawn from these exciting results, however, it needs to be established that this star is indeed representative of how discrete absorption components are formed, at least in some O stars. In other words, is ξ Per unique?

Finally, the decay observed in both γ Cas and ξ Per seems to suggest that the scatter in the "memories" of these two stars is real and

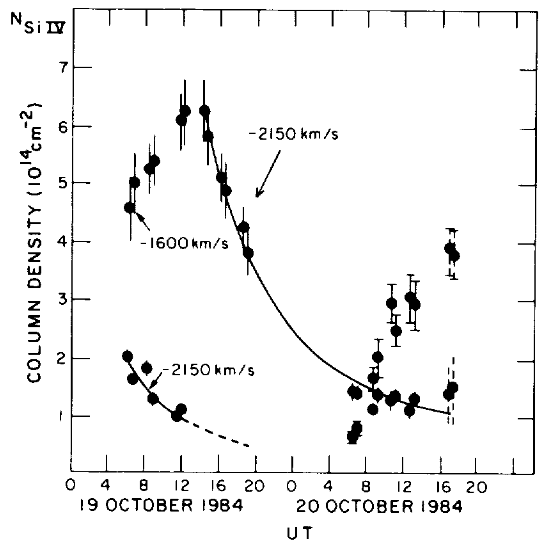


Figure 4-45. Column-density measurements of the October 1984 spectra of ξ Per. The lower left group of points corresponds to the high-velocity component visible in the first few spectra of October 19. The central group pictures the column density of the broad and, later, narrow absorption components, initially accelerating with about 35 km/s per hour. A $(t_0 + t)^{-2}$ fit is superposed. The lower right group (shown with different formal error-bar shapes) is the beginning of the new cycle.

depends on the particular epoch during an episode in which the spectrum is taken.

4. Discussion. Detailed studies of a few stars (γ Cas, ξ Per, 59 Cyg, and ω Ori) have shown that, in our opinion, the appearance of discrete components occurs on a short time scale (hours) and that the subsequent decay takes much longer (days or weeks). If this were the case for all stars with discrete absorptions, the fact that, in some stars, the features appear to be more or less permanently present (but variable) can be explained by assuming a rather frequent production of "new" narrow lines superimposed on the "old" ones, as occasionally observed in ξ Per. Detailed studies of more stars are clearly needed to verify or deny this extrapolation. The study of 66 Oph by Barker and Marlborough (1985) probably did not have high enough time resolution to establish the short time scales.

The important issue of uniqueness of discrete component behavior is ξ Per can be settled only when more such studies are made. The fact that ξ Per showed a memory similar to γ Cas, ω Ori, and 10 Lac is suggestive of a common origin in both O and Be stars. Presently available data in IUE archives are not of a sufficiently high time resolution to answer this in the affirmative.

A third important point is concerned with the uniqueness of the line-fitting method. The rapidly changing complex Si IV profiles in ξ Per (Figure 4-44) can be fitted equally well with a continuously changing "underlying" P Cygni profile with additional absorption (see Prinja et al., 1987). It is not clear, however, how to interpret such a fit because the method used to calculate P Cygni profiles assumes a steady-state (time-independent) configuration, which is obviously not the case. It is clear that many more model calculations are necessary before we may conclude what information about the geometrical structure the observed profiles actually reflect. (See also Section E.4 for a further discussion of model calculations.)

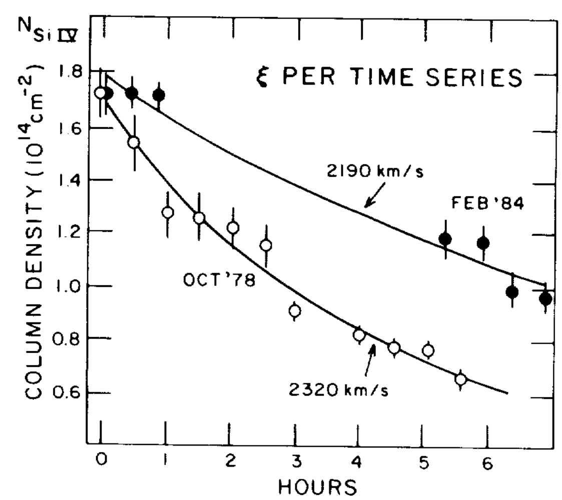


Figure 4-46. Decaying column density of Si IV narrow absorptions on two earlier occasions observed in ξ Per (Prinja et al., 1984). The velocities at which they occurred remained the same during the decay, but were significantly different.

Finally, in our opinion, the longer time scale mentioned previously reflects changes in the star itself, rather than a time scale determined by the flow properties of a stellar wind. Growth of instabilities, for instance, takes place on a time scale much shorter than the flow time scale.

E. Models

In a spectral line indicating mass outflow, there are several ways to form a discrete high-velocity absorption component:

(a) The presence of a velocity plateau (or a decrease in velocity gradient) which will increase the path length of absorbing ions at a given velocity.

(b) The existence of an ion density that is higher at a specific velocity than at neighboring velocities. This might be caused either by an overall enhancement of the density at that velocity or by ionization effects which favor the observed ionization fraction at that velocity.

(c) As described in terms of the model by Lucy and White (1980) and Lucy (1982a, 1982b, 1983), an increase of either the number of scattering surfaces or the shock amplitudes in a flow with a non-monotonic velocity law will increase the number of resonances of a photon and, hence, the optical depth at a given velocity.

All of the proposed models for the *formation* of a discrete absorption feature fall in one of these categories. A basic unresolved question about their *origin* is: Given the ubiquity of the absorption features, do they reflect the intrinsic behavior of any stellar-wind flow or are they caused by the underlying star in a way that we observe the resulting interaction with the stellar wind? We shall see that, from the observational side, the major constraint imposed on a model does not come from the formation of discrete components, but from the predicted time variability. Earlier reviews and discussions of current models are given by Lamers et al. (1982), Howarth (1984), Henrichs (1984), and Prinja and Howarth (1986). In the following, we summarize and briefly comment on the different models which can be considered for explaining the behavior and origin of narrow absorptions. Many of these models were proposed without the present knowledge of the time variability.

(a) Peaks in the radial ionization structure in the stellar wind: Excluded by the lack of correlation of velocity with ionization potential. In addition, a plausible mechanism for changing the ionization balance appropriately does not exist, especially regarding the high-velocity features which cannot be formed close to the photosphere.

(b) A stationary shell in the wind: Excluded by the observed variability.

(c) Velocity plateau(s) in the stellar-wind flow (e.g., Hamann, 1980): Doubtful in view of the frequent occurrence of multiple components and variability. Detailed model calculations by Prinja and Howarth (1984) of the effects of a velocity plateau on theoretical P Cygni profiles also showed that velocity plateaus, given the assumed properties, are unlikely to provide the correct explanation.

(d) Postcoronal decelerated region of low temperature (Doazan in Underhill and Doazan, 1982, p. 394) explaining the absence of absorption at intermediate velocities by the existence of a steep gradient in the ionization balance. The deceleration comes from two interacting mass flows. Difficulties similar to those in (a), (b), and (c) apply.

(e) A stellar wind decelerated by the gravity of the star or by running into the interstellar medium: Found unlikely by Lamers et al. (1982).

(f) A two-component stellar wind (high and low density) caused by the inherent instabilities of a radiation pressure driven flow (Lamers et al., 1982). (Various instabilities have been discussed by Lucy and White (1980), Carlberg (1980), Kahn (1981), Lucy (1982a, 1982b, 1983, 1984), Owocki and Rybicki (1984, 1985), and Krolik and Raymond (1985).) A model like this (based on time-averaged flow properties) is intrinsically unable to describe the coherent time behavior of discrete absorptions as observed in some stars (γ Cas and ξ Per), although, of course, the predicted effects of clumpiness, the production of X rays, etc. in a flow with a nonmonotonic velocity law remain unaltered. At present, this possibility has not been applied to the aspect of time-dependence. In general, such analyses can predict a time scale for the development of an instability, which turns out to always be much shorter than the flow time scale, the latter apparently being the characteristic observed time scale of variability. Therefore, the flow instabilities might be related to the narrow absorption components, but an additional ingredient is needed to explain the observed time scales. Lucy (1983) mentioned the formation of an additional high-velocity absorption component in a stellar wind as a result of a shadowing effect caused by multiple shocks. This remark, however, is primarily concerned with the origin of the high-velocity layer(s) and not as much with its precise location or time behavior. Further investigation is worthwhile.

(g) Temporary release of parcels of matter from "open" magnetic loops which are attached to, but above, the stellar surface, and consequently coupled with rotation (Underhill and Fahey, 1984): Difficulties in explaining the absence of predicted strong, low-velocity, suddenly disappearing absorption components in addition to ad hoc assumptions about geometry and time of ejection of parcels. (See Section E.1 for a more detailed discussion.)

(h) Corotating interacting regions (CIR's) arising from (assumedly) coexisting high- and low-velocity regions in the stellar wind which, at their interface, may produce a velocity plateau, analogous to that observed in the solar wind. In this model, proposed by Mullan (1984), rotation is also an essential ingredient. In the way CIR's are formed in the description given by Mullan, however, the fast streams are identified with the terminal velocity of the wind, which presumably carries the bulk of the matter, opposite to the solar case. Such an assumption makes it difficult to explain the often encountered saturated P Cygni profiles in O stars, but cannot be excluded in Be stars. In addition, an expected correlation between the ratio $v_{rot}/v_w(\infty)$ and the column density of discrete components can not be found, with the present data, among O stars and B supergiants, but could not be tested for Be stars. For these reasons, apart from the necessary (but probably not unreasonable) ad hoc assumptions about the existence of coexisting high- and low-velocity material, we consider it unlikely that the concept of CIR's in its present form can be applied to explain the occurrence of discrete absorptions in the O and luminous B stars, but might find some application in Be stars. (See Section E.2 for a more detailed discussion.)

(i) Expansion of a high-density layer in the stellar wind, proposed by Henrichs et al. (1980), elaborated in Henrichs et al. (1983) and Henrichs (1984): This descriptive model is probably the most simple one for explaining or predicting the time-dependent behavior of the discrete components as observed in some stars, but it gives no clue of the cause of the origin of the high-density layer. (Episodic mass loss has been proposed.) This model was rejected by Lamers et al. (1982) because, at that time, no substantial short time variations had been observed.

In the following, we give more detailed discussions of the three last-mentioned models, as they enable specific predictions regarding the occurrence and time behavior of a discrete absorption component which can be compared with observational data. It should be made very clear in advance, however, that by no means does a "final" picture exist or that the correct model is even included in any of them; these models hopefully provide a good starting point for future research. Considering the wide variety of behavior of discrete absorptions, it is more likely that ingredients of more than one of the foregoing models might apply. For instance, the assumption of spherical symmetry in some of these models is certainly not correct in the case of the rapidly rotating stars, but it is unknown how large the deviations are.

1. The Model of Underhill and Fahey. To explain the existence and behavior of discrete components, Underhill and Fahey (1984) envisage that early-type stars have localized spots above (but not on) the stellar surface from which parcels of gas are released in addition to a uniformly emitted steady stellar wind. They argue that such local spots can arise only from magnetic field configurations on the surface of a star and therefore postulate the existence of "closed" and "open" magnetic loops, the latter being regions in which the magnetic field lines extend far into space and from which the parcels are released. The configuration resembles that of solar corona models. It is beyond our scope to discuss those assumptions concerning the magnetic field configurations associated with early-type stars, but we will consider the predictions for the behavior of discrete absorptions implied by this model. Figure 4-47 illustrates the model. A star with radius R_* rotating with angular velocity ω_0 emits a parcel

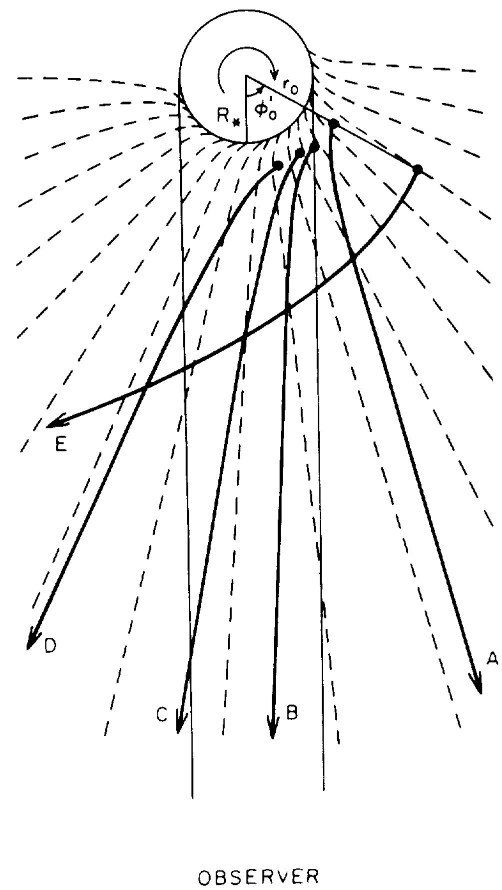


Figure 4-47. Stellar-wind trajectories (dashed lines) and "parcel" trajectories (heavy solid lines) in an inertial frame illustrating the model of Underhill and Fahey (1984). Angle ϕ_0' represents 360° - ϕ_0 in the terminology of Underhill and Fahey. The figure is drawn to scale for the equatorial plane of a star with $v_{rot}/v_w(\infty)$ = 360/1700 (resembling ζ Oph) and a velocity law with steepness β = ½. An observer will never see an absorption component from parcel A, but will from parcels B, C, D, and E. Parcel B will cause a relatively long-lived absorption feature. Parcels C, D, and E will each produce shorter living low-velocity features. The model predicts an equal likelihood for production of parcels A, B, C, and D. The observational absence of strong, suddenly disappearing absorption features might lead to the conclusion that this model cannot represent the behavior of the discrete components.

of gas from location $\vec{r_o}$, $\vec{\theta_o}$, $\vec{\phi_o}$, which will subsequently follow a trajectory r, θ, ϕ, in a plane containing r_o and $\omega_o \times r_o$, thereby conserving its angular momentum. Underhill and Fahey present expressions for the trajectories of both wind particles (originating at R_*) and parcels (originating at r_o in an inertial frame). To facilitate the discussion, we consider only parcels in the stellar equatorial plane (i.e., $\theta = \theta_o = 90°$, as drawn in Figure 4-47), where we note that, for different colatitudes, the 0 dependence will be reduced by a projection factor. (See Underhill and Fahey, 1984.) We will discuss the velocities, lifetimes, and column-density behavior, respectively.

The velocity of a parcel $v_p(r)$ can be readily calculated from the wind velocity $v_w(r)$ by making the (reasonable) assumptions that the parcel acceleration is the same as that of the wind and that the rotational energy can be neglected (as $v_{rot}{}^2 << v_w^2(\infty)$). This implies that

$$v_p^2(r) - v_p^2(r_o) = v_w^2(r) - v_w^2(r_o) \quad . \quad (4\text{-}42)$$

It is obvious that the terminal velocity of a parcel will always be less than the terminal velocity of the wind if the initial velocity of the parcel is smaller than the wind velocity at the location where the parcel was emitted. By identifying the observed radial velocity of a discrete component with the approximate terminal velocity of a parcel, Underhill and Fahey (1984) infer that "average" parcels must be emitted from about $r_o = 2\ R_*$, depending on what velocity law is assumed. The closer the point of release is to the star, the less the difference between $v_p(\infty)$ and $v_w(\infty)$. We conclude that the predicted velocity range is in agreement with the observations.

However, the observability of a parcel as an absorption component strongly depends on the point of origin of the parcel. This can be readily seen from Figure 4-47, which is drawn to scale for a star with $v_{rot}/v_w(\infty) = 360/1700$ (it is this ratio that determines the pattern of wind and parcel trajectories), representative of a star like ζ Oph, O9.5 V(e). (For stars with a smaller ratio, $v_{rot}/v_w(\infty)$, the pattern would deviate less from purely radial motion than in Figure 4-47, but the following reasoning would remain unchanged.) A parcel following trajectory A has no chance to cross the line of sight to produce an observable absorption component. A similar parcel emitted at the same distance from the star but a little later (B) is obviously capable of causing an observable absorption feature. A parcel (C) emitted still later crosses the column of sight only for a while, causing a strong absorption feature at low velocity which will accelerate and suddenly disappear when the parcel trajectory leaves the line of sight. Parcels following trajectory D will cause strong, very short-lived absorption features at low velocity. If emitted at the appropriate time in the stellar rotation cycle, a parcel emitted from a larger distance from the star (e.g., E) may stay in the line of sight, but will enter it with a higher fraction of its (lower) terminal velocity. From this geometry, it is clear that an observable absorption component can be produced only by a parcel which has the correct initial location and which is launched within a time window that is rather small with respect to the stellar rotation period.

To be compatible with the observational fact that, for the majority of the stars, the discrete absorptions appear to be present for most of the time (if not always) and are variable in both strength and velocity (the latter to a lesser extent), the Underhill and Fahey (1984) model requires that, for a given star, parcels are allowed to be released only from approximately the same spot relative to the observer (i.e., *only at the same phase in the stellar rotation cycle*). This is because, if they were emitted from other longitudes as well, one would expect cases C and D (stronger, suddenly disappearing lines) to prevail over the longer living lines (case B), in strong disagreement with the observations. However, if we accept the existence of such a preferred spot above the stellar surface (we cannot exclude, however, parcels being released from different colatitudes because they will not be observable), we find it unlikely that so many stars show discrete components. On the other hand, a continuous stream of matter with a higher density than the ambient wind being released from a single spot can immediately be excluded because, in such a "lawn-sprinkler" model, multiple absorption components will always be present, each of which arises from the different gas layers crossing the line of sight at equidistant radial intervals, causing a new component to appear during each rotation cycle. Such a strictly periodic behavior has not been observed. We can also exclude intermediate cases, in which only a small fraction of a sprinkler pattern exists because of a stream of matter released from one point for a time period shorter than the stellar rotation period. We would expect to see as many trailing edges (as in cases C and D) as leading edges. In conclusion, it is the absence of suddenly disappearing strong absorption components that presents a major difficulty in the Underhill and Fahey model.

Another concern in this model is the size of a parcel at its release and consequent expansion due to geometric and pressure effects. The fact that there are many reported cases of essentially saturated discrete absorption lines (see for example, Prinja and Howarth, 1986, and Figure 4-35) means that the material must cover most if not all of the stellar disk as seen by the observer. This also seems to be implied by the observed enhanced emission. (See Figure 4-44.) In such cases, relatively small areas, as pictured by Underhill and Fahey (1984), are not appropriate.

A final point we want to mention is the time dependence of the column density as predicted by the Underhill and Fahey model. After a parcel has expanded to cover the full stellar disk, its column density will rapidly decrease according to $(t_0 + t)^{-2}$, where t_0 is a constant. To account for the observed, rather constant range in column density (Figure 4-35), new parcels have to be released rather frequently to replenish the absorption. As stated previously, the only way to accomplish such a replenishment is to repeat the release of parcels at instants which are separated in time by an integer multiple of the stellar rotation period, which required an unusual ejection mechanism. We conclude that the Underhill and Fahey model is unlikely to be compatible with the behavior of discrete absorption components as observed in most stars. However, it might have its application in some Be stars, among which such a wide variety of different behavior has been observed. (See for instance, the examples in Figure 4-34.) On the other hand, the high flexibility of the Underhill and Fahey model makes it difficult to make specific predictions as to precisely what to expect in such cases. See Bates and Halliwell (1986) for additional discussion and extended application of this model.

2. Corotating Interacting Regions. Mullan (1984) proposed that, if fast and slow streams coexist in a stellar wind arising from different longitudes on the stellar surface, the fast stream will catch up with the slow stream when the stellar rotation is taken into account. The strongly shocked interaction region between the two streams might cause a velocity plateau and an enhanced density region (called a CIR), which would then be responsible for a high-velocity discrete absorption component. The picture closely resembles the behavior of the solar wind as observed near the ecliptic plane by means of space probes.

A first point of objection is related to the proposed applicability of the analogy between the Sun and hot stars. Mullan (1984) identifies the velocity of fast streams in OB stars with the terminal velocity of the wind. In many O stars, $v_w(\infty)$ is derived for a given star from saturated P Cygni profiles, whereas the discrete components are observed in unsaturated lines. This makes one think that the terminal velocity of the wind is reached over most of a 4π solid angle seen from the star and will carry the bulk of the outflowing matter. This is in contrast to the solar case, in which the fast streams are believed to originate in the much smaller coronal holes.

A second point is concerned with the rotation rate of the stars. In Mullan's (1984) description, the wind speed is a function of longitude at the stellar surface. Of course, this function is unknown, but would probably show up as a rotationally modulated stellar wind. No clear indications for such an effect have been found thus far. On the other hand, when, in terms of this model, the assumption is made that the wind speed becomes different beyond a given distance from the stellar surface, the model resembles more and more the model of Underhill and Fahey (1984) (see the previous section) with similar implications.

Another point is that the CIR model predicts that the ratio $v_{rot}/v_w(\infty)$ controls the radial distance at which the discrete components begin to form. Because one might reasonably assume that the associated column density will decrease as a function of distance, a correlation between the ratio previously mentioned and the column density is expected for a randomly chosen collection of stars. When we use the sample of Prinja and Howarth (1986), we find that a possible trend (if significant) for the N V data appears to be opposite to that of C IV, when $V \sin i/v_w(\infty)$ is plotted against the column density of the discrete components. This is difficult to explain.

At the moment, it appears unlikely that the CIR model in its present form is capable of explaining the observed properties of discrete components in O stars and B supergiants. Like the model of Underhill and Fahey (1984), it might find its application in some Be stars, where we do not directly observe the high-speed wind in the asymmetric line profiles and where spherical symmetry is absent. At present, not enough data exist to support such a hypothesis. Additional discussion on this model (from a different point of view) appears in Prinja and Howarth (1986).

3. Expansion of a High-Density Layer. This descriptive model was originally called the "UV-shell model" because it considers the ultraviolet discrete absorption lines which are observable only in lines of ions which are stellar-wind indicators; these are situated mostly in the UV. Figure 4-48 (from Henrichs, 1984) displays the characteristics of the model. At a certain epoch, the star produces a high-density layer in its wind. The origin of this behavior is not specified; it might be caused either indirectly by enhanced mass loss during a short time (hours), spherical (shells) or nonspherical (puffs), as proposed by Henrichs et al. (1983), alternatively by instabilities in the flow as

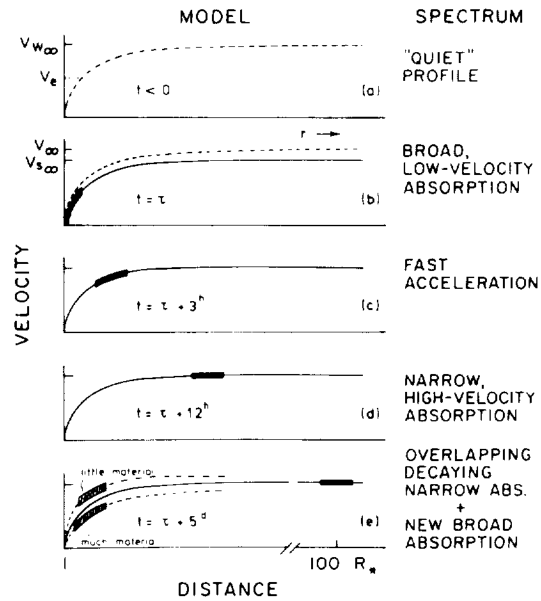


Figure 4-48. Schematic representation of the time history of a high-density layer (causing discrete absorption components), spherically expanding in a stellar wind (Henrichs, 1984). (a) A typical velocity profile of an accelerating wind (dashed line). $v_{w\infty}$ denotes the terminal velocity of the wind, which is always observed. (b) When a high-density layer (caused by enhanced mass loss, by instabilities in the flow, a combination thereof, or by some other process) begins at some epoch τ at low velocity, a broad absorption feature will be observed. The acceleration of this layer is slower than that of the continuous stellar wind (dashed line) because the density is higher. (c) The time scale for acceleration is of the order of hours. (d) Within 12 hours, a narrow high-velocity feature will be observed. Its strength will decay simply because the layer geometrically expands. (e) A possible explanation for a memory as observed in some stars: when the layer contains much material, it might undergo less acceleration than when there is little material. Multiple structure will be observed when more layers coexist.

discussed in model (f), by a combination of the two (e.g., an instability triggered by enhanced mass loss), or by some other (unknown) process. The rapid expansion of this high-density region is presumably due to the same acceleration mechanism that accelerates the stellar wind: radiation pressure (Lucy and Solomon, 1970; Castor et al., 1975a; Abbott (1980, 1982a). Because of the higher density of this layer, its "terminal" velocity is expected to be equal to or smaller than the terminal velocity of the continuous wind. How much smaller will depend on how effectively saturation reduces the radiative acceleration. This might be a very qualitative explanation for the "memory" of some stars; when the layer contains much material, its terminal velocity will be less than when the layer contains less material. A quantitative comparison with model calculations and observations has yet to be carried out. As mentioned previously, the short acceleration time (likely to be different for different stars) explains why the discrete absorption lines are observed mainly at high velocity. The faster the acceleration, the smaller the probability of finding them at low velocity. In conclusion, the overall picture in terms of this model is that, when only one high-density layer is involved, one expects to see the well-known $(t_0 + t)^{-2}$ behavior of the column density as soon as the velocity is at its maximum value (assuming constant ionization fractions). The associated lifetime is of the order of days to weeks. (γ Cas might be such a case.) When more layers, which are unlikely to have exactly the same properties, are involved, they will cause overlapping absorption features, each with its own time constant, giving the impression of a more or less constant, but variable, feature at high velocity. This picture might be applicable to many O and B stars. When the time resolution of the observations is short compared to the replenishing time scale, one might see the acceleration, as seen for instance in ξ Per, Figure 4-44.

We note that, in this model, as in almost any model of a high-density layer that has a significantly lower velocity than the terminal velocity of the ambient stellar wind, this layer must be strongly shocked toward the star, where the (less dense) wind runs into the backside of the layer. This might be the reason for the observed ionization differences between the P Cygni profile and the discrete components as found by Lamers et al. (1982) and Prinja and Howarth (1986).

Arguments against this model for stars other than those mentioned previously are the absence of low-velocity material in some of the best monitored cases, like ζ Oph, and the sometimes very irregular behavior in some Be stars, like 66 Oph. (See for example, Figure 4-34g, based on Grady et al. (1984a).)

Finally, it should be emphasized that, in this model, the predicted time behavior follows directly from simple geometric effects in which the expansion speed dictates the observed time-dependence. In other words, no physical effects such as ionization, etc. are taken into account.

4. Discussion. Simple model calculations by Prinja and Howarth (1985) showed that a high-density layer caused by episodic mass loss, and accelerating at the same rate as the ambient wind, would not cause an absorption feature until it reached a substantial fraction of its terminal velocity, making the probability of finding low-velocity absorption even smaller (Figure 4-49a). The same spherical model calculations fit reasonably well with the few existing observations of the narrowing during the acceleration phase. Prinja and Howarth's model calculations also showed that rather frequent ejections of "shells" are necessary to maintain a "long-lived" more or less stationary absorption feature (Figure 4-49b). On the other hand, those calculations are based on a simple enhancement of the mass-flux parameter, keeping all other flow properties unchanged. This is certainly not a realistic assumption because it neglects the interaction of the stellar radiation with the matter in the layer and, hence, ignores a probably different acceleration. In addition, the ionization structure is imposed in those models, whereas in reality, it is actually a function of the radiation field and shock

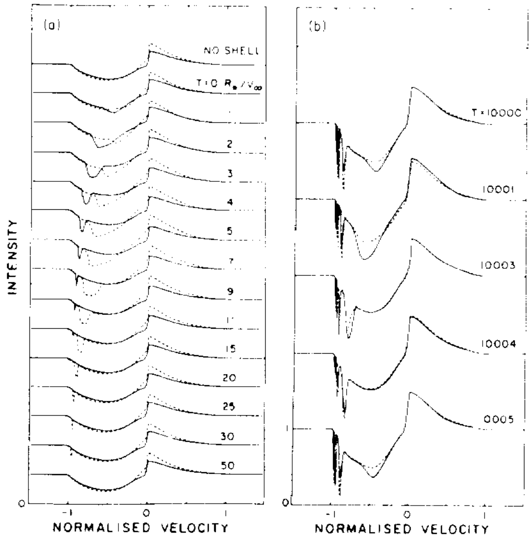


Figure 4-49. Model calculations by Prinja and Howarth (1985) of the effects of an enhanced mass loss on a P Cygni profile. (a) Absorption due to a single shell as a function of time (in units of $R_*/v_w(\infty)$) does not begin to show up in the profile as a discrete feature until it reaches about one half of the terminal velocity. Solid and dotted lines denote a velocity law with $\beta = ½$ and 1, respectively. (b) Multiple shell ejection with time intervals of 5 $R_*/v_w(\infty)$ for $\beta = ½$. Dashed lines denote results for an arbitrary density enhancement factor which is an increasing function of velocity.

properties. For instance, the positive slope in the memory (Section D.2) as found in HD 199579 (Figure 4-42), if significant, cannot be explained by this model without invoking a change in the run of the ionization balance with distance. Such a change would also alter the expected statistics of occurrence of low- and high-velocity material. A detailed model which considers those effects has not been constructed. Nevertheless, it is encouraging to see that at least some of the observed properties of the discrete components as observed in some stars can be fairly represented.

Another issue is the question of spherical symmetry. Within the scope of current theoretical considerations, expanding spherical shells are expected to give rise to decaying emission (at the red side), eventually disappearing beyond detectability. The emission phase is expected to be much shorter than the absorption phase, but should nevertheless be observable. Because there are only a few reports about enhanced emission, a tentative conclusion is that the line-forming regions are probably not spherically symmetric. However, the fact that a large fraction of stars show discrete absorptions suggests that the material subtends a significant solid angle as seen from the star.

In the model described above, the intriguing question remains: What causes the high-density region(s) in the wind? Is it a basic property caused by the star itself or does it find its origin in the flow? Many model calculations must be carried out before we can answer those questions. Note that multiple shocks in a stellar wind, as proposed by Lucy (1982a) are believed to be primarily responsible for the production of X rays throughout the wind. X-ray emission from early-type stars, however, appears to be always present (e.g., Cassinelli et al., 1981; Seward and Chlebowski, 1982) and sometimes variable (e.g., Snow et al., 1981), but the variability does not seem to be as dramatic as in the discrete components. Hence, if multiple shocks are indeed the cause of most of the observed X-ray emission, they seem unlikely, at first, to be the cause of the much more variable discrete absorption lines. One might argue, however, that we are dealing with X-ray emission integrated over a volume and, in the case of discrete absorptions, only with line-of-sight effects.

Simultaneous observations in other wavelength bands are also helpful. We give a short account of a few of them which are, in our opinion, significant.

F. Related Observations

If a high-density region in a stellar wind would originate in the flow, there would be a minimum velocity at which a discrete absorption can be observed. However, different physical conditions that exist close to the star (temperature, density, and ionization) might also prevent observation of the low-velocity features in the UV. (Compare the model calculations of Prinja and Howarth (1985).)

Significant in this respect, therefore, might be the transient absorption features observed in the Lyman δ and ϵ lines in several OB stars by Gry et al. (1984) with the Copernicus satellite. Figure 4-50 gives an example of α Vir, B1 IV, in which low-velocity absorption components (-80 and -140 km/s) disappeared within 8 hours. Gry et al. found a similar

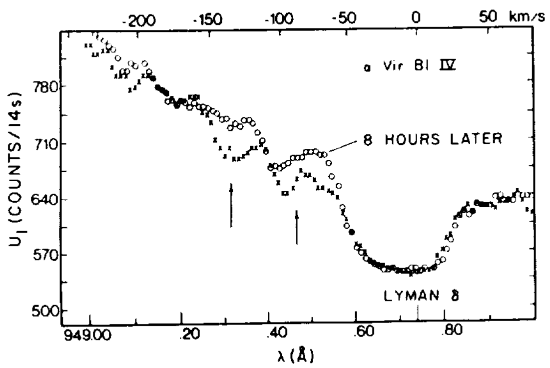


Figure 4-50. Transient absorption features in Lyman δ in α Vir observed with Copernicus by Gry et. al. (1984). The time scale is indicated. It has been suggested that these features precede the occurrence of the high-velocity narrow absorptions.

behavior in a major fraction of their sample (5 of 8 stars). The authors concluded from these observations that an expanding high-density shell, more or less similar to the model described previously, could fit the observed behavior and that these lines might be the precursors of the high-velocity narrow absorptions as described in this review. In addition, they tentatively suggested that the phenomenon may be related to an instability of the outer envelope of the star, similar to the well-known instability proposed for the β Cep (or β CMa) stars, which are situated in the HR diagram between spectral types B0 and B2. (See for example, Lesh, 1982.) It is well known, however, that the region in the HR diagram where the discrete components occur is much wider; moreover, a physical relation between the Be phenomenon and the β Cep stars has never been proposed. However, variable nonradial pulsation, which many investigators now consider to be the origin of the variable behavior of many stars in this region of the HR diagram, is attractive as a possible mechanism for stellar-wind variability.

1. Nonradial Pulsations. Evidence that the study of nonradial pulsations might help us to understand the origin of the discrete absorptions has been suggested (Henrichs, 1984) on the basis of existing (accidentally) simultaneous optical and UV observations of the O9.5 Ve star, ζ Oph. Vogt and Penrod (1983) presented persuasive arguments that the observed transient distortions of the He I λ6678 line in this star can be explained by high-order nonradial pulsations. Furthermore, they found a clear correlation between the observed amplitude of the oscillations from season to season with the well-documented outbursts (as seen in Hα) of this star. Vogt and Penrod suggested that "sporadic release of pulsation energy during mode switching might provide a plausible explanation for these outbursts and for the occurrence of the Be phenomenon in general." Henrichs (1984) noted that the drastic changes in the discrete absorption components in IUE spectra of ζ Oph found by Howarth et al. (1984) actually occurred more or less simultaneously with the changes in amplitude of pulsations reported by Vogt and Penrod. This implies that structural changes in the outer layers (the subatmosphere) of the star, in this case seen as changes in nonradial pulsation behavior, are coupled to structural changes far out in the wind, seen as varying discrete absorptions. Similar cases (as found by Henrichs, 1986) are λ Eri, B2 IIIe, and 6 Cep, B2.5 Ve, for which Barker and Marlborough (1985) described drastic changes in the UV C IV profiles and noticed their covariability with Hα-emission activity, whereas simultaneous changes in the nonradial pulsation behavior had been reported independently by Smith and

Penrod (1985). (See also Penrod (1986).) A montage of Hα, C IV, and nonradial pulsation behavior in He I λ6678 (after Penrod, 1986, and Barker and Marlborough, 1985) is shown in Figure 4-51. Baade (1985) discusses implications of the widely encountered variable nonradial pulsations for modulating mass fluxes in Be stars. See also his review about O stars in the present volume.

2. Nonradial Pulsations and Stellar Winds. Given that in ζ Oph the pulsational amplitude (peak to peak) is about 6 percent in stellar radius and 21 km/s in velocity with a period in the range of 7 to 15 hours (Vogt and Penrod, 1983), an interesting question is whether there can be a smooth, continuous, isotropic wind under such conditions. The location of, and the physical conditions at, the sonic point (or its analog in the case of radiation pressure (Abbott, 1980)), play a crucial role in any supersonic flow. The base of the wind (i.e., on the surface of the star) moves up and down over different parts of the surface with a velocity

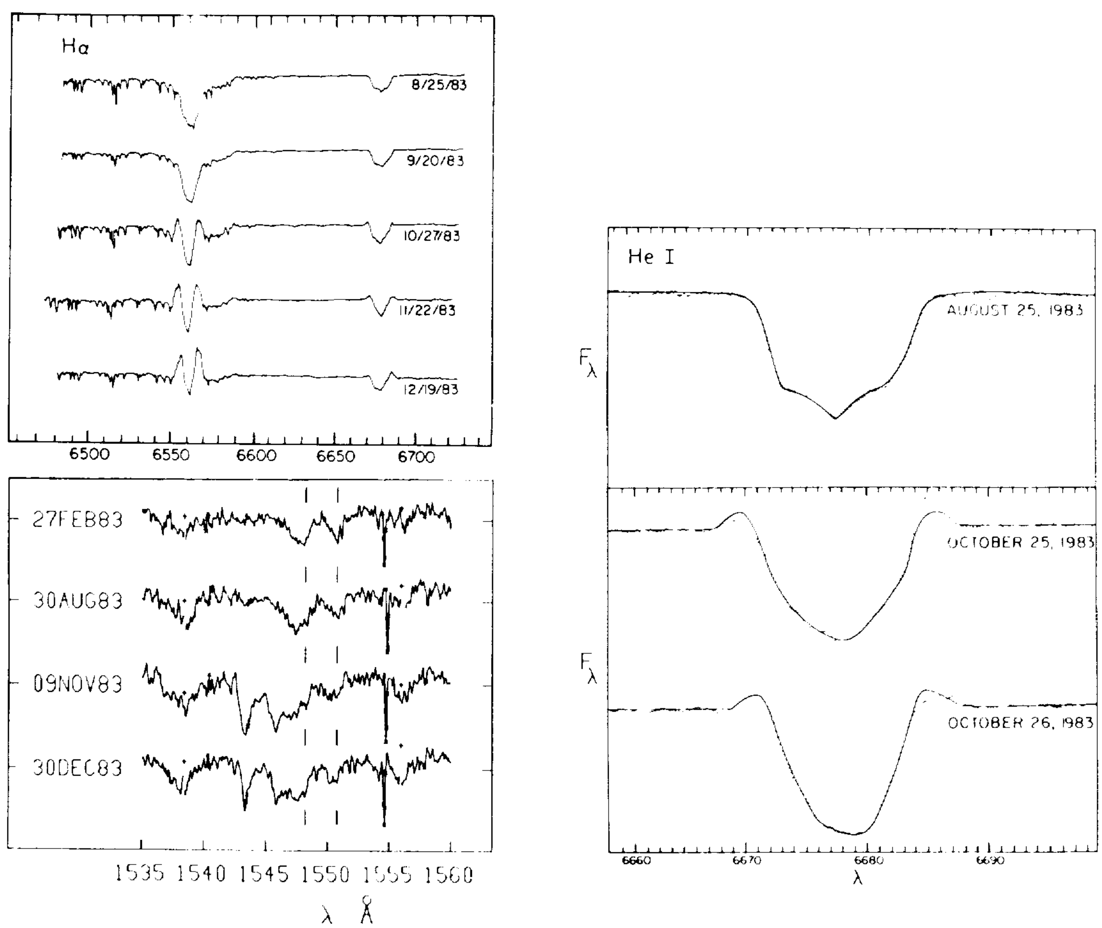


Figure 4-51. A montage of (unplanned) approximately simultaneous Hα, C IV, and nonradial pulsation behavior in He I in the B2 IIIe star, λ Eri (after Penrod, 1986, and Barker and Marlborough, 1985). Note that, when Hα began to develop emission, discrete components in C IV absorption appeared (rest wavelengths are indicated). According to Penrod, the amplitudes of the nonradial ℓ = 2 and ℓ = 8 modes decreased *roughly a factor of 2 in the same period (observed profile shown as dots, theoretical fit shown as solid line). These observations suggest that the energy lost in pulsation might be directly transferred to the stellar envelope.*

comparable to the local speed of sound. The region where the wind becomes supersonic must have a very irregular as well as time-dependent shape, making a stable continuous flow highly unlikely. Some theoretical considerations by Castor (1986b) indicate that a shock front will form below where the sonic point would be in the wind of a nonpulsating star. If the pulsation period of the star is longer than the typical flow time scale (roughly 10 to 1 in the case of ζ Oph), the wind will be relatively insensitive to the motion of the photosphere, but the detailed effects on the flow are still subject to considerable study. (See the references cited in Section E.)

A second point to be made is a remark on Vogt and Penrods' (1983) suggestion of mode-switching as a cause of release of pulsational energy. The origin of switching between different oscillatory g modes is not known. (See for example, the reviews by Unno et al. (1979), J. P. Cox (1980), and A. N. Cox (1983).) The point is that pulsational energy being liberated in the outer layers during such a switch might trigger enhanced mass loss for which, unfortunately, no quantitative calculation exists. The occasionally reported examples of mode-switching sometimes appear to be questioned (Penrod, private communication), but variations in pulsation amplitude seem to occur rather often. These are also associated with changes in pulsational energy (of the typical order of 10^{44} erg), but whether and how such energy can be deposited in the stellar wind is not clear. Finally, it should perhaps be stressed again that it is the *variability* in the pulsation behavior of the star that is proposed to be connected with the formation of discrete components. For other issues related to pulsation and stellar-wind behavior, see papers referred to in Abbott et al. (1986).

3. Discussion. There are additional attractive features for suggesting future investigations of a possible connection between amplitude-changing and/or mode-switching of nonradial pulsating stars and the occurrence of discrete components. First, spectroscopic studies indicate that the presence of nonradial pulsations in many O, B, and Be stars is well established. (See for example, Bolton (1982), Baade (1983, 1984a, 1984b, and private communication), Smith and Penrod (1985), and Penrod (1986).) In addition, the well-known irregular photometric behavior of almost all luminous early-type stars (e.g. Maeder, 1980b, and de Jager, 1980) can be explained by (nonradial) pulsational behavior. (See for example, the discussion by Cox (1983, p. 146).) Of course, the uniqueness of such interpretation is a serious matter of discussion. (See for instance, Baade's chapter in this book.) From a nonradial pulsation point of view, Penrod (1986) asserts that the difference between Be and B stars is that most Be stars pulsate in a long-period $\ell = 2$ mode and usually in a short-period high mode as well, whereas "normal" B stars pulsate in one or two short-period $\ell = 4-10$ modes. The apparent existence of this difference in energy reservoir of these two types of stars is another interesting indication that might help us to understand the variability in Hα emission and its relation to the more frequent UV outbursts (here identified with discrete absorption-line episodes) which do not occur in the low-luminosity non-Be stars. (See Section B.) Extension to the hotter O stars, however, has yet to be established. Significant in this respect is that Baade (1985, and this volume) has found clear evidence for nonradial pulsations in ζ Pup, O4 I(n)f, for which a detailed record of its discrete-component behavior exists (Prinja, 1984). Simultaneous optical and UV observations are clearly needed. At this point, a speculative remark regarding the role of rotation in Be stars can be made. It is well known (in theory) that the presence of rotation allows both nonradial and toroidal modes to be excited and also widens the frequency range in which a star can pulsate, as compared to a similar but nonrotating star. (See any textbook.) It is therefore suggestive to speculate that variability in pulsational behavior is a more common phenomenon among the rapid rotators simply because it seems "easier" to excite a certain mode in a star with a wider range of frequencies and modes, thus marking the significance of rapid rotation to the Be phenomenon. As long as no excitation mechanism is known, however, such a point of view is difficult to prove.

Another point is that the time scales of nonradial pulsation are of the same order (hours) as the flow time scales, which can therefore cause drastic changes in the base of the flow of the wind which might trigger enhanced mass loss and/or instability in the flow as discussed previously. Furthermore, the observed but unexplained irregular behavior of velocity-amplitude changes and/or mode-switching of many nonradially pulsating stars follows a pattern similar to the transient activity of the discrete components in some stars. The longer time scale, mentioned before, might be identified with the long time scale involved in changes in pulsation behavior.

The nonspherical symmetry induced by introducing nonradial pulsations as an additional mechanism to trigger or regulate mass loss is another favorable aspect. In rotating stars, the amplitude of nonradial pulsations with modes $\ell = |m|$ is strongly peaked toward the equatorial plane, where the gravity is also reduced, hence favoring a preference plane. From the observational side, particularly from the polarization and infrared data, it seems unavoidable that, in the case of Be stars, the mass flux is not spherically symmetric. A link between those two configurations must be established, however, before conclusions can be drawn. For an additional discussion, see Stalio and Zirker (1985).

Note that one aspect of the description given above in some sense approaches the view expressed by Thomas (e.g., 1973, 1982, 1983, and many other papers) that the mass flux of a star is, in fact, a parameter which does not depend only on the radiative properties of its atmosphere, in contrast to the approach of the pure radiatively driven mass-flux theory. Baade (1985) describes in detail a working hypothesis of how variable nonradial pulsations can modulate the mass flux in Be stars and how the stellar-wind structure will be influenced by such effects. Other discussions about nonradiative activity in hot stars can be found in Underhill and Michalitsianos (1985).

G. Concluding Remarks

We wish to add a few points to the previous issues. First, the very important related question of superionization in stellar winds and the possible association with X-ray generation must be mentioned. (See for example, Odegard and Cassinelli, 1982, and Marlborough and Peters, 1982.) For example, much of the superionized C IV absorption in most Be stars appears to exist in the form of high-velocity features. On the other hand, intrinsic X-ray emission from Be stars has been studied only fragmentarily (e.g., Peters, 1982a), and any possible connection between them is unknown. Some aspects of this problem might be studied by comparing the time-dependent behavior of variable discrete absorption lines arising from different ionization stages in connection with X-ray studies. Another point worth mentioning is the possible connection between UV discrete absorption activity and the long-term Be-Bshell-B sequence as observed in many Be stars (e.g., 59 Cyg and γ Cas; e.g., Doazan et al., 1983; Doazan et al., 1985, and Doazan and Thomas, 1986).

Also of clear importance are simultaneous studies in different wavelength bands, which include polarization measurements, as carried out for instance for ω Ori by Sonneborn et al. (1987). Intrinsic polarization variability in O and B stars on time scales of days to months appears to be common (Lupie and Nordsieck, 1987). Simultaneous UV and polarization modeling might give some clue to the degree of spherical symmetry and/or latitudinal dependence of stellar winds (e.g., Brown and Henrichs, 1987). Such latitudinal effects are most likely to occur in the case of Be stars. (See for example, the discussion on μ Cen by Peters in Stalio and Zirker (1985). Binarity is another possible means for studying the narrow line phenomenon. Howarth (1984) did not find any change in the velocity of the discrete absorptions in the B0 supergiant, V861 Sco, in spite of the clear binary motion of the star. McCluskey and Kondo (1981) note that, in the interacting binary system, AO Cas (consisting of two type O stars), the discrete components definitely vary, but that their data were not with sufficient time resolution to establish whether or not a binary orbital phase exists. It might be difficult to untangle the secular from orbital effects. Narrow lines are also found in at least one X-ray binary (4U0900-40, Sadakane et al., 1985) containing a B supergiant and an X-ray pulsar, opening the possibility of probing the stellar-wind structure by observation of pulse-period changes (van der Klis and Bonnet-Bidaud, 1984). (For a review of the available data and theory on the response of the rotation rate of a neutron star accreting from a stellar wind, see Henrichs (1983).)

Also of interest are the observations by White et al. (1983), who reported rapid X-ray spectral variability from the X-ray binary, 4U1700-37, which was interpreted as being caused by small-scale inhomogeneities in the stellar wind of its optical counterpart, HD 153919, O6.5 Iaf. Unfortunately, the C IV, N V, and Si IV lines in the UV are completely saturated in this star, but similar studies in other wind-fed massive X-ray binaries might be more successful.

Finally, we want to emphasize that one has just begun to realize that variability is a basic signature of stellar winds of all early-type stars and that understanding this phenomenon is essential for a complete understanding of stellar-wind structure. In this chapter, we have tried to focus on the observational evidence for wind variability and have reviewed suggested interpretations and possible related issues. "The best possible data set" to prove or disprove our present thoughts has yet to be constructed. It is the author's hope that this review might help to reach this goal.

H. Acknowledgments

This review is an extension of the one given at the Fourth European IUE Conference, Rome, Italy, May 1984. I am grateful for the discussions on several aspects of this review with Dave Abbott, Paul Barker, Joe Cassinelli, Eugene Damen, Tony Hearn, Ian Howarth, David Hummer, Henny Lamers, Keith MacGregor, Derck Massa, Dean Pesnell, Raman Prinja, and many participants of the workshop, "The Connection Between Nonradial Pulsations and Stellar Winds in Massive Stars," held in Boulder, Colorado, April 1985, all of whom contributed much to shape my thoughts. I would like to thank especially Ian Howarth and Raman Prinja for permission to quote their results prior to publication and in general for the very pleasant collaboration over the years. I thank David Hummer for a detailed reading of the manuscript and, together with Dave Abbott, Derck Massa, Raman Prinja, Anne Underhill, and especially Dick Thomas for constructive remarks on earlier drafts. Thanks also to the Regional Data Analysis Facility (RDAF) in Boulder (supported by NASA contract number NAS5-26409), where the skillful and friendly help of Terry Armitage has been indispensable. Finally, I want to thank Peter Conti for his optimistic patience in waiting for this manuscript and for his generous hospitality at JILA. Part of this research was financially supported by the Niels Stensen Foundation, which is gratefully acknowledged.

Note added in proof: After this review had been completed, new observations have shown that the reoccurrence time scale of discrete-component episodes in O stars might well be correlated with the stellar rotation period (cf. Figure 4-44b).

5

ENVIRONMENTS AND EVOLUTION

I. ENVIRONMENTS OF STARS

A. Conceptual Framework

The term "interstellar medium" (ISM) has come to be applied to galactic material not already collected into stars or other gravitationally bound objects. It can include dust clouds and molecular clouds, and it is known to be highly heterogeneous in its properties. We know that the overall energetics in the ISM lead to a concept of a "three-component" model for the gas (McKee and Ostriker, 1977). These include a low-density hot ($\sim 10^6$ K) region which cannot cool easily, a warm ($\sim 10^4$ K) volume, and the cool ($\sim 10^2$ K) high-density clouds. According to McKee and Ostriker, the hot region takes up most of the space in the galactic disk; warm regions surround the considerably smaller cold interstellar clouds which are more massive. In the conventional picture, supernova explosions are the dominant energy source to the hot low-density regions (Cox and Smith, 1974). Abbott (1982) has suggested that O-star winds may also inject a similar amount of energy and may dominate in regions of OB associations (see also Van Buren, 1985). The term "violent interstellar medium" (McCray and Snow, 1979) describes the complicated physical phenomena occurring inside a galaxy such as ours.

The observations of the interstellar "extinction," or reddening, are associated with a dust or "grain" component toward stars distant from the Sun and the presence of absorption lines with the gaseous material along the lines of sight. Some of this component near the very luminous O and Wolf-Rayet stars appears to be modified (e.g., ionized, excited, or dissociated) by the star's Lyman continuum radiation and stellar winds. I believe it would be useful to describe this modified ISM by a new term—local stellar medium (LSM).* The name then immediately calls attention to the fact that this material, both the dust and the gas, may have been altered from the preexisting ISM state by the star.

Circumstellar material has sometimes been used as a phase to describe the medium I would prefer to call LSM. However, the term "circumstellar" connotes gas and/or dust which has been ejected from a star and remains near or even bound to it. The LSM may have a component of ejecta from the star, but the distinction here is that the bulk of the material is preexisting ISM. The LSM term is useful if it can be identified with an observed star or stars; it probably is not needed in cases in which the ISM is modified and distorted by such events as supernova explosions nearby.

O and Wolf-Rayet stars form preferentially (entirely?) in large star clouds called associations. Considerable gas and dust is associated

*This has been called the "local stellar environment" by others (e.g., Doazan and Thomas, 1983; Thomas, 1983).

with these star-forming clouds. The ISM in our Galaxy is very clumpy and is not uniform along a line of sight. We might expect that, for relatively nearby O and Wolf-Rayet stars, the LSM contribution to the overall extinction and to the line-absorption features might play a potentially significant role. For example, if there are three clouds on the average per Kpc in the galactic plane (Shull and Van Steenberg, 1985), then an O or Wolf-Rayet star at 1 Kpc will probably be born and remain in one of the three. The modification to the birth cloud, or preexisting ISM, by the star itself will thereby possibly affect the observed reddening and interstellar line contribution seen along the line of sight.

The central regions of the Orion nebula, near the Trapezium, would be a good example of material which could be readily described as the local stellar medium rather than an interstellar medium. This gas and dust is under considerable interaction from the O star's ultraviolet radiation and from the stellar winds, resulting in large-scale turbulent and chaotic motions, a high ionization state, and modified dust particles. The "reddening law" in the Orion region is well known to be anomalous.

Other obvious evidence of an LSM can be found in observations of wisps of nebulosity or more structured "rings" associated mainly with Wolf-Rayet stars. These are most readily seen in direct imaging in narrow filters centered on the O III line at λ5007 Å or Hα (e.g., Chu et al., 1983).

In other cases, the LSM can be easily identified in direct photometry of extensive and extended regions surrounding the OB associations themselves, as in the extreme case of 30 Dor in the LMC. Here, the action of many O and Wolf-Rayet stars creates and completely dominates the "giant H II" region which surrounds them. Very complex physical phenomena are occurring, many features of which are not well understood.

This general topic is one to which an entire book could be devoted. In this chapter, I will touch briefly on hot-star environments to call the attention of the reader to perhaps extreme examples of nonthermal phenomena associated with the O and Wolf-Rayet stars that are the topic of this monograph.

B. Theoretical Considerations and Models

Castor et al. (1975b) were among the first to discuss the interactions of stellar winds with the ambient interstellar medium (see also Steigman et al., 1975). Their use of the term "bubble" has been adopted in the literature to describe the consequences of the dynamic forces at work. The generalized but highly simplified theory and the mathematical details are given by Weaver et al. (1977). The following simplifying assumptions were initially made: (1) the ambient interstellar medium is homogeneous in density; (2) the wind is spherically symmetric and steady; and (3) the star is at rest with respect to its environment. These authors identified four separate volumes surrounding a star with a stellar wind. In each of these volumes, different physical phenomena play leading roles. Good synopses of the generalized concepts are given by McCray (1983), Lamers (1983), and Kwok (1984).

In the innermost region (a), the stellar wind flows hypersonically, having previously pushed out the preexisting material. The time for the initial free expansion is of the order of 10^2 years. At first, the size of this region is some few tenths of a parsec, depending on the total amount of swept-up material and the strength of the wind. Inside this volume, the normal wind equations (e.g., Abbott, 1986) hold. This region continues to expand, more slowly after the initial rapid phase, eventually reaching dimensions of the order of a few parsecs. (One parsec is some 10^6 R_* for main-sequence O stars.)

After the very rapid initial wind expansion, an adiabatic phase begins. In this region (b), there is a hot (10^6 K) volume of shocked stellar wind with a small fraction of swept-up interstellar material. The kinetic energy of the stellar wind is deposited into this *coronal* region, which itself is also continuing to expand in what is called the *snowplow phase*. The expansion is described as "energy conserving" or "momentum conserving," depending on the radiative losses in the region. This hot bubble should emit coronal lines and soft X rays, but the latter would be hard to detect against the general diffuse background. The size of the bubble depends on the wind strength of the central star and the density of the ambient ISM, but for typical conditions, it is of the order of a few to tens of parsecs.

Surrounding the hot bubble is a relatively dense shell (c) of swept-up interstellar material. This shell is the result of the slow expansion of the coronal region. Most of the mass of the bubble is found in this volume, which will have a temperature of some 10^4 K. Some of the shell mass diffuses back into region b. Here, the density may be sufficient that *high-excitation* interstellar lines such as O VI are seen (Weaver et al., 1977), although these features can also be found at the interfaces between the supernova shells and the cooler gas in the ISM (Jenkins, 1978a, 1978b; McCray and Snow, 1979).

The bubble gradually evolves over time, expanding against the ambient ISM outside of shell c. The lifetime of the bubble is a few 10^6 yr, set by the evolution time of the central exciting star. The thick shell (c) may provide a contribution to the *low-excitation* interstellar lines (McCray and Snow, 1979). Figure 5-1 provides a schematic view of a typical bubble from an O-type star after 10^6 yr (expanding into a uniform ISM of density 1 cm^{-3}; adapted from Weaver et al., 1977). Much more mass is swept up by the expansion of the bubble than is provided for in loss by the star. In the terms I am using here, the LSM encompasses volumes a, b, and c.

More complicated situations clearly exist in the real ISM. Shull (1980) has treated the case in which the star and its stellar wind are embedded in a molecular cloud, which is a probable initial condition. If the initial ambient medium is very inhomogeneous, the coronal volume b may contain slowly evaporating dust clouds. These would act as *cooling centers* and alter the snowplow expansion phase from an

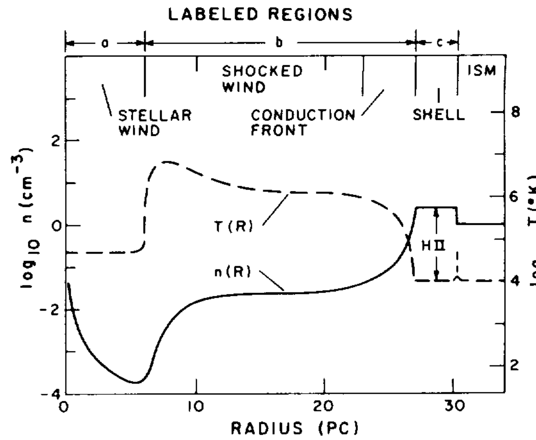


Figure 5-1. Schematic large-scale temperature/density structure of an idealized stellar-wind bubble (adapted from Weaver et al., 1977), expanding into a uniform ISM of density $M_\odot$ = 1 cm^{-3} at 10^6 yr. The wind parameters are v_{TERM} = 2000 km s^{-1} and $\dot{M}$ = 10^{-6} $M_\odot$ yr^{-1}. For a typical O star, the ionized H II region is at a radius some three times that of region c.

energy-conserving case to a momentum-conserving one (McCray, 1986). The physics of this situation has not yet been completely worked out. In another example of complications in the *real world*, if the ISM density is sufficiently large and the Lyman flux insufficient (as from a typical late O-type or early B star), the H II region may be trapped inside region c. This would result in a large density discontinuity at the outer boundary region, with possible observable consequences (Hollenbock et al., 1976). McKee et al. (1984) have considered the idealized case of stellar-wind bubbles expanding into a cloudy ISM. Shull (1982) has given a generalized summary of some of the dynamic effects of stellar winds and shocked gas on the interstellar clouds.

C. Observations

Stromgren spheres, or ionized hydrogen (H II) regions, surround all stars that emit Lyman continuum radiation. The treatment of the relevant equations is relatively straightforward for a homogeneous ISM, but in nearly all observed cases, inhomogeneities are present. The H II volumes can be detected by free-free radio emission and the recombination spectrum of hydrogen (radio and visible) and other common ions, such as O III, O II, and N II. Direct imaging of excited regions of space can be made with narrow-passband filters centered on appropriate lines, such as Hα and O III. Examples of this technique are given in the survey of the galactic plane by Parker et al. (1979). Nebulosities associated with H II regions can be found in their photographs.

There are many instances of local stellar material being directly observed and associated with single O and Wolf-Rayet stars. The so-called *ring nebulae* (Johnson and Hogg, 1965; Smith, 1967) have been loosely defined as arcs of nebulosity centered on and ionized by Wolf-Rayet stars; occasionally other objects appear to be the exciting stars (e.g., HD 148937 and NGC 6164-65). I will touch briefly on a few of the issues concerning ring nebulae. Chu and her associates, in an important series of papers (references in Chu, 1982), have made a systematic and detailed study of the optical properties of the ring nebulae associated with Wolf-Rayet stars. Using strict selection criteria, they list 15 known in our Galaxy and nine in the Large Magellanic Cloud (LMC).

Chu (1982) points out that one may expect the interaction between the central Wolf-Rayet star and its environment to be of three types: stellar UV radiation, stellar-wind ejecta, and the wind bubbles. She identifies and labels four types of ring nebulae, based on the image morphology and kinematic data: structured (R_s) and amorphous (R_a) H II regions, stellar ejecta (E), and windblown bubble (W). These correspond, respectively, to the expected classes. The three categories should not be considered as mutually exclusive because all Wolf-Rayet stars have strong UV radiation. One half of the galactic-ring nebulae can be identified as being in the bubble phase. The classifications of the LMC rings are more difficult to make unambiguously, partly because of the greater distances away from the Earth. Chu suggests that the ring morphology compared to the spectral type of the central star implies that WC-star environments are more advanced than those of WN stars. The R-type nebulae appear to be relatively old (10^6 years); the W nebulae are younger (10^4 to 10^5 years).

Although the more careful searches of Chu and her associates and Heckathorn et al. (1982) have turned up a few more rings than were previously known, not all Wolf-Rayet stars have them. Chu suggests that this is because many isolated stars are now located in the hot low-density ISM where the conditions are not right for ring formation or ease of detection. In many other cases, the existing Wolf-Rayet stars are found in crowded regions in which O stars still reside; the effects of the winds are therefore much more complicated.

The W-type rings around single Wolf-Rayet stars appear to have roughly the physical characteristics predicted for them by the Weaver et al. (1977) theory. The situation is complicated by the past evolutionary history of the Wolf-Rayet star, and the very heterogeneous nature of the surrounding ISM. The E rings may represent material ejected from the object that is the progenitor of the Wolf-Rayet star (e.g., in a red-supergiant stage (McCray, 1983)). The R rings merely represent the interaction of the radiation from the star with the LSM.

The spectra of seven ring nebulae surrounding WN stars have been analyzed by Barker (1978), Kwitter (1981, 1984), and Peimbert et al. (1978). In nearly all cases, the nitrogen/helium ratio is enhanced compared to the normal. The analysis is relatively model-insensitive since two ionization stages of oxygen are present, thus determining the excitation conditions in the nebulae. The interpretation is complicated by the fact that the ring is composed of stellar-wind material and swept-up ISM. Kwitter shows that one can calculate the fractional contribution of the wind by two methods: the first compares the nebular ion abundances to those of the star itself; the other estimate comes from the dynamic age of the nebula and the mass-loss rate of the central star. In most cases, these fractions (of the order of a few percent) agree to within the errors, thus suggesting consistency in the abundance determinations and the concept of windblown bubbles. In one discrepant case, Kwitter believes that the ring is due primarily to an ejection *event* rather than to a windblown bubble (in agreement with Chu's (1982) morphological classifications).

Several WC stars may have rings associated with them (Lortet et al., 1982), but a detailed study of one such possible association (Lortet et al., 1984) reveals complex morphology and nearby O stars as possible contributors. Kwitter (1984) has pointed out the importance of abundance analyses for isolated WC-ring nebulae if such could be found.

Some WC stars have dust *shells* surrounding them (e.g., Cohen et al., 1975). These are typically detected from excess emission at a few micrometers with a characteristic wavelength dependence indicating relatively cool temperatures of the order of 1000 K. Nearly all low-excitation (WC9) stars show the presence of the dust signature; in most cases, there appears to be little variability from year to year, implying relatively constant rates of dust formation in the stellar winds (Williams, van der Hucht, and Thé, in preparation). The shell has a size of some $10^4\ R_*$ in the case of the WC9 star, WR 104, for which speckle IR photometry resolved the emission (Dyck et al., 1984).

Variable IR emission and changing conditions in the inferred dust shells have been noted in a few other WC stars: HD 193793 WC7 (Williams et al., 1978; Hackwell et al., 1979); HD 192641 WC7 (Williams et al., 1985); Anon WC9 (Danks et al., 1983). The former two stars have had episodic changes in the IR emission, implying changing modes of dust formation. These two bright stars have also shown visible emission-line variations which suggest changing mass fluxes, but the data are insufficient to permit correlation of changes in the different wavelengths with each other. The two WC7 stars also have hydrogen absorption lines in their spectra from a companion star, but the relationship of these features to the variable dust emission is not understood.

The formation of relatively cool dust in the winds of WC stars is not very well understood. Presumably, the dust shells are composed of graphite because the "silicate" absorption feature near 10 micrometers is not present (Cohen et al., 1975). The enhanced carbon abundance of WC stars may be a necessary condition for the formation of dust, but this is purely speculative at present. Hackwell et al. (1979) suggest that iron particles may be responsible for the variable shell surrounding HD 193793. Williams et al. (1985) suggest that the episodic emission in HD 192641 comes from a radius of approximately $2 \times 10^3\ R_*$, somewhat smaller than those of the WC9 stars. Because the stellar winds vary in some of these objects, the dust formation may be enhanced or suppressed accordingly. Far more data in many spectral regions will be needed to help our understanding of these curious phenomena.

II. MASSIVE-STAR EVOLUTION

According to the current "standard" theory, main-sequence evolution can be conveniently divided into three subgroups according to mass (e.g., de Loore, 1980). These divisions are for low-mass stars ($M < 1.2\ M_\odot$), intermediate masses ($1.2 < M < 10\ M_\odot$), and massive objects ($M > 10\ M_\odot$) according to the characteristics of core behavior during core hydrogen burning and the subsequent evolution in following stages. We are concerned here with Population I O and Wolf-Rayet stars whose initial masses are $>10\ M_\odot$ and $>40\ M_\odot$, respectively. Massive stars end their lives as supernovae of type II, then as neutron stars or black holes; the low- and intermediate-mass objects may all end as white dwarfs, although in cases of close binary interactions, some intermediate-mass objects may also evolve to the neutron state (possibly by the type I supernova mechanism).

How massive are the most massive stars? Stability considerations (Schwarzchild and Harm, 1958) suggested 60 $M_\odot$ as an upper limit, but Appenzeller (1970a, 1970b) and others pointed out that nonlinear theory could, in principle, raise this limit to 100 $M_{\odot}$ or more. There is very suggestive evidence, based on locations of the most luminous stars in the HR diagram, that the largest masses are somewhat over 100 $M_{\odot}$ (Humphreys and Davidson, 1979). The central objects in 30 Dor may have masses up to 200 $M_{\odot}$ but are otherwise similar to those known in the field (e.g., Walborn, 1986). In this section, then, we must consider the evolution of star masses from 20 to 200 $M_{\odot}$. I wish to briefly consider the recent theoretical work in this area because it directly affects our understanding of the relationship of Wolf-Rayet to O-type stars.

P Cygni lines have long been known to be present in Wolf-Rayet stars. With their discovery in the resonant UV lines in normal OB supergiants (Morton, 1967) and the inference of mass loss by stellar winds, considerable attention has been paid to the consequences for stellar evolution models. Chiosi and Nasi (1974) were among the first to consider a set of models with mass loss specifically included. One difficulty with this work, and nearly all that has followed, is that the mass-loss rate has to be parameterized. Approaches have used either a theoretical expectation (e.g., following Castor et al.'s (1975b) formula) or empirical determinations (e.g., Lamers, 1981). Because there is thus far no thorough understanding of the physical basis of mass loss, these theoretical models can be only approximations of the real situation. They are useful, however, and although not the final word, they point us inexorably to the importance of stellar winds in massive-star evolution. The details will certainly change as better models and improved mass-loss rates are developed.

A number of groups have been constructing evolutionary models of single massive stars with mass loss included. These are Chiosi and associates (e.g., Chiosi et al., 1978, 1979), de Loore and associates (e.g., de Loore et al., 1977, 1978a, 1978b), Brunish and Truran (1982a, 1982b), Maeder (1981, 1983), and Stothers and Chin (1980). These models showed broad similarities in the sense that the end point of core hydrogen burning was moved to cooler effective temperatures, fainter bolometric magnitudes, and longer MS lifetimes than the object would have had in the absence of mass loss. Depending on the mass-loss rate adopted (and the parameterizations adopted by the various authors often differed substantially), either very little or considerable mass was lost by the end of core hydrogen burning.

In progressively more massive stars, the initial convective cores are increasingly larger fractions of the total stellar mass. If sufficient mass is lost that the initial convective core is uncovered, the star will show an enhanced surface helium content. The subsequent evolution will then be back toward higher effective temperatures, and no red-supergiant phase will occur.

The more massive stars are observed to have larger mass-loss rates because they are brighter. It turned out that, except for the very brightest stars, none of the rates were quite high enough to reach the initial convective core before the end of core hydrogen burning. However, almost all of them were of sufficient magnitude. Noels et al. (1980) suggested that the adoption of slightly larger rates could lead in a natural way to the convective core being uncovered in the most massive stars ($M > 80$ $M_{\odot}$) during core hydrogen burning, and these were not excluded by the observations. Noels and Gabriel (1981) presented more detailed calculations with this possibility included and followed the evolution of all light elements, not just hydrogen and helium, on stellar surfaces.

With the construction of models with mass loss, it was realized that the observed locations of massive stars in the theoretical luminosity versus effective temperature plane held clues to the evolution. There are basically two distinct characteristic anomalies in the comparison of observed and theoretical HR diagrams. Humphreys (1978, 1979) and Humphreys and Davidson (1979) showed that surveys of luminous stars in the Local Group galaxies revealed the existence of blue supergiants which are considerably brighter in bolometric magnitudes than the red supergiants. Thus, above some M_{bol} which did not depend on the galaxy type or its composition, no red supergiants existed, but blue supergiants were found. Because massive stars evolve at roughly constant bolometric magnitude, this meant that the helium-burning phase, normally associated with the red supergiants, was occurring elsewhere. Secondly, compared to their core O-star progenitors, there were too many B supergiants. The number of the latter stars should be only 10 percent of the former, given the location of the termination of the core hydrogen-burning tracks in the HR diagram (Stothers and Chin, 1977; Chiosi et al., 1978; Cloutman and Whitaker, 1980; Meylan and Maeder, 1982). This feature has been referred to as "main-sequence widening."

To illustrate these considerations, Figure 5-2 (adapted from Conti, 1984) shows a plot of those galactic luminous stars within 2.5 kpc on an HR diagram, along with tracks from Maeder (1983). There is almost a two-magnitude difference in the luminosities of the brightest blue and red supergiants. Note also that, between the tracks for 30 and 60 $M_\odot$, there are many stars to the right of the termination of the core hydrogen-burning tracks. Thus, the two problems: massive stars burning helium do not appear as red supergiants, and the tracks are inadequate to explain the main-sequence widening.

Maeder (1983) suggested scenarios in which mass loss in the following shell hydrogen-burning phase (or core hydrogen-burning phases), which he identified with Hubble-Sandage variables, could result in sufficient mass being lost that helium-enriched material appears on the stellar surface of stars with masses as low as 60 $M_\odot$. These models also then turn leftward in the HR diagram. In this paper, he gives details of chemical abundance changes as the stellar models evolve. Maeder also pointed out the likelihood that the most massive stars would evolve quasi-homologously (i.e., chemically mixed).

Bressan et al. (1981) pointed out that *overshooting* from the convective core, not previously accounted for, might be an important ingredient in the models. The idea is as follows. At the core boundary, the thermal gradient leads to a boundary condition such that the mass acceleration is zero. The material, however, could well have a velocity which would carry it above the boundary, thus overshooting the nominal core. Thus, the composition of material outside the core might partially be contaminated from that inside. In that case, the initial convective core in a massive star could be uncovered with less mass loss than would be needed without this effect. Bertelli et al. (1984) and Doom (1982a, 1982b) have presented series of models which incorporate these effects. Maeder (1982) has constructed models with overshooting and has pointed out the possibility that turbulent diffusion might play a role in massive-star evolution.

In all the models used thus far, the overshooting parameterization is arbitrary because the physical situation is not well understood. Various authors adopt different philosophies for including overshooting and must necessarily have a mass-loss parameterization already explicitly included. Thus, they obtain different predictions as to the effect of the inclusion of overshooting on the models. For example, Doom (1982a, 1982b) uses as a criterion for overshooting the condition that below some mass a star will become a red supergiant, and above that mass it will turn back toward higher temperatures. He finds that the bifurcation mass is 33 $M_\odot$ with his models when he compares them to the Humphreys and Davidson's (1979) limit. Prantzos et al. (1986) have recently published a new set of evolutionary calculations for massive stars. Their bifurcation mass is 42 $M_\odot$, and they note that the value depends on the mass-loss rate, among other parameters not completely determined.

It appears to me that both mass loss and overshooting are very important ingredients in the evolution and dramatically affect massive stars. The details are subject to revision as better data and theory are incorporated. Although

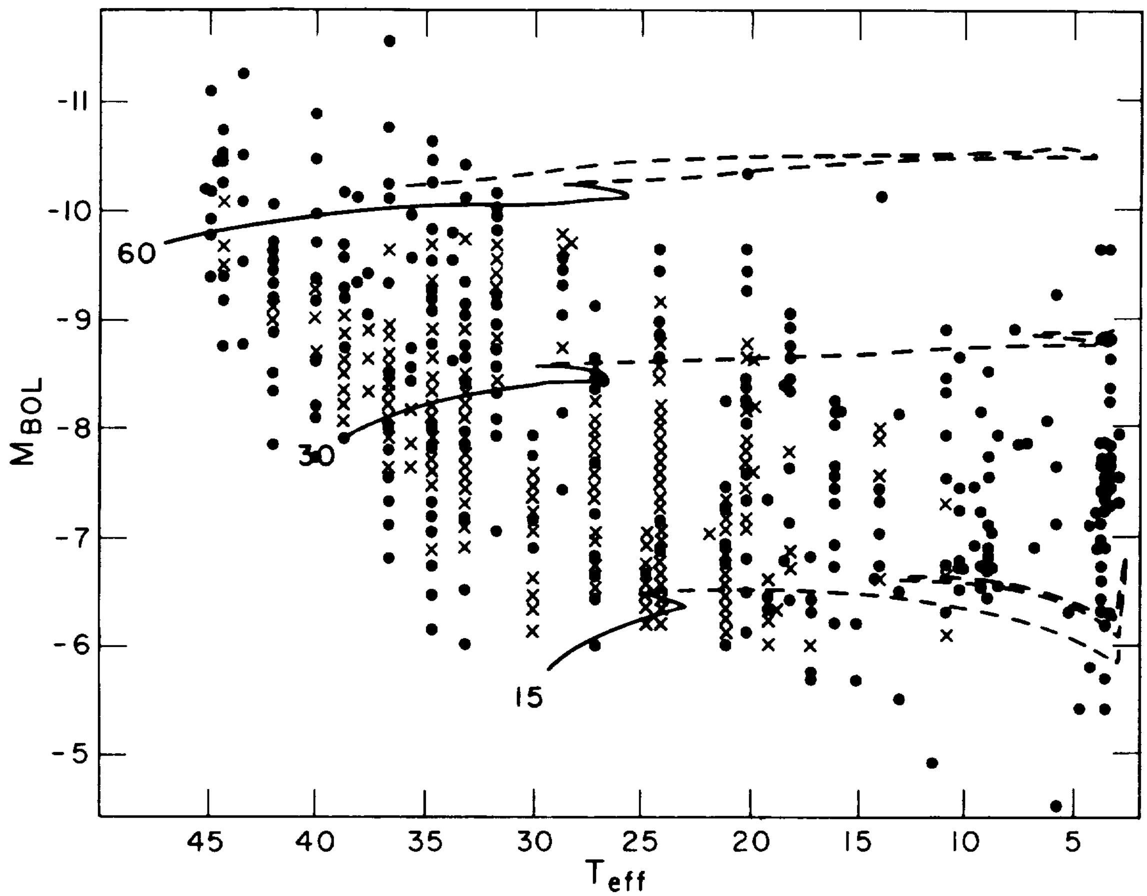


Figure 5-2. Observational/theoretical HR diagram for luminous stars in our Galaxy within 25 kpc from the Sun (from Conti, 1984). The filled symbols denote single stars; the crosses denote more than one star at that position, up to 10 or so. The evolutionary tracks are those of Maeder (1983).

the final details are not yet included, the general importance of mass loss and mixing is recognized. Doom et al. (1986) give an update of their calculations and evolutionary scenarios for massive stars. A thorough review of issues of massive-star evolution has recently been given by Chiosi and Maeder (1986). The interested reader can find many more details and complete references in these papers.

III. EVOLUTIONARY SCENARIOS FOR MASSIVE STARS

We have now assembled all the ingredients for a discussion of potential connections between massive O stars and the Wolf-Rayet objects. Let me briefly summarize the *facts* as presented in many of the previous pages.

The stars with O spectral type are among the hottest and are the most luminous of all stars. Their effective temperatures range from 30000 to 50000 K, and their masses range from 20 $M_\odot$ to perhaps ten times that. These stars are all to be associated with the core hydrogen-burning stage, which may extend to late B supergiants when the effects of *main-sequence widening* are considered. The stars are tightly confined to the galactic plane and represent well the objects of extreme Population I. The most massive stars are also preferentially found in the inner spiral arms of the Galaxy in comparison with objects in the outer Perseus arm.

Within factors of 2, most O stars have normal composition, although those in the Magellanic Clouds, particularly in the Small Magellanic Cloud, may well be deficient in CNO and higher atomic-number elements. No red-supergiant stage is observed to correspond to or represent the helium-burning phase of the most luminous O stars.

The Population I Wolf-Rayet stars are also very luminous, with effective temperatures from 30000 to 90000 K and masses from 10 to 40 $M_\odot$. These stars are also tightly confined to the plane of the Galaxy, and like the massive O stars, they are preferentially found in the spiral arms interior to the Sun. Their compositions are not normal but show evidence of enhanced helium and nitrogen in the WN types or overabundant helium, carbon, and oxygen in the WC types. The actual anomalies are qualitatively consistent, with the WN stars showing the products of CNO equilibrium hydrogen burning and the WC objects showing the results of helium burning. The Wolf-Rayet stars are observed to be overluminous for their mass, which is to be expected for helium-burning objects.

One straightforward interpretation of these results is that the Wolf-Rayet objects represent the helium-burning phase of massive-star evolution, which is not otherwise identified with observed stars. This follows from their apparent composition anomalies, their overluminosity for their mass, and their similar galactic distribution to the massive O-star population. They must be descendant from the O stars because their masses are lower and their abundances are abnormal.

The spectra of Population I Wolf-Rayet stars represent the extremes of hot mass-losing objects in which the wind is sufficiently strong that emission lines dominate in the visible region. These stars are potentially hot enough and sufficiently luminous that the winds can be accelerated by purely radiative forces. On the other hand, variability and nonradial pulsations may play a potentially important role in determining the mass-loss rates and the structure of the wind. The ionization balance requires nonthermal energy input of a more extreme nature than is found in the similar situation for O- and Of-star winds.

The ionization sequence of the WN and WC subclasses is not well understood. Whether it represents an effective temperature sequence is not clear, nor is its relationship to evolutionary constraints. Those WN stars with hydrogen would be considered less evolved than those without. Hydrogen is detected preferentially among the later types but occurs in all WN subclasses. Although it appears clear that the WC stars are more highly evolved than the WN classes, given their composition, there is no clear pathway among the various subtypes.

The hot luminous stars of types O and Wolf-Rayet represent natural physical laboratories in which extreme conditions of high radiation pressure, nonthermal processes, and departures from equilibrium are found. As such, they can potentially lead to better understanding of complex physical phenomena not otherwise amenable to description. There is now considerably better knowledge of some of their overall features than existed a few years ago, and observations and accompanying theoretical understanding and improved modeling are proceeding apace. It is indeed an exciting time to be studying and investigating these rare and exotic objects. Observations with the Space Telescope and other satellite instruments should greatly increase the available data; improved modeling of the stellar interiors, atmospheres, and winds, along with the boundary regions in the interstellar medium, will be a necessary adjunct in the future.

I appreciate discussions concerning O and Wolf-Rayet stars with Drs. de Loore and Massey. I began this manuscript while I was a Visiting Astronomer in residence at Cerro Tololo InterAmerican Observatory. I am indebted to the then Director, P. Osmer, and the staff there for an extremely pleasant and productive visit. This work has been supported by the National Science Foundation, most recently under grant AST85-20728.

REFERENCES

Abbott, D.C. 1978, *Astrophys. J.*, **225**, 893.

Abbott, D.C. 1979, in *Proc. IAU Symp. 83, Mass Loss and Evolution of O-Type Stars*, ed. P.S. Conti and C.W.H. de Loore, p. 237.

Abbott, D.C. 1980, *Astrophys. J.*, **242**, 1183.

Abbott, D.C. 1982a, *Astrophys. J.*, **259**, 282.

Abbott, D.C. 1982b, *Astrophys. J.*, **263**, 723.

Abbott, D.C. 1985, in *Relations Between Chromospheric-Coronal Heating and Mass Loss in Stars*, 1984 Workshop, Sacramento Peak Observatory, Sunspot, New Mexico, ed. R. Stalio and J.B. Zirker, p. 265.

Abbott, D.C. 1986, in *Proc. Third Trièste Workshop, Relations Between Chromospheric-Coronal Heating and Mass Loss in Stars*, ed. R. Stalio and J. Zirker (Trièste: Tabographics).

Abbott, D.C., Bieging, J.H., and Churchwell, E. 1981, *Astrophys. J.*, **250**, 645.

Abott, D.C., Bieging, J.H., and Churchwell, E. 1984, *Astrophys. J.*, **280**, 671.

Abbott, D.C., Bieging, J.H., and Churchwell, E. 1985a, in *The Origin of Nonradiative Heating/Momentum in Hot Stars*, ed. A.B. Underhill and A.G. Michalitsianos, NASA CP-2358, p. 47.

Abbott, D.C., Bieging, J.H., and Churchwell, E. 1985b, in *Radio Stars*, ed. R.M. Hjellming and D.M. Gibson (Dordrecht: Reidel), p. 219.

Abbott, D.C., Beiging, J.H., Churchwell, E., and Cassinelli, J.P. 1980, *Astrophys. J.*, **238**, 196.

Abbott, D.C., Bieging, J.H., Churchwell, E., and Torres, A.V. 1986, *Astrophys. J.*, **303**, 239.

Abbott, D.C., Bohlin, R.C., and Savage, B.D., 1982, *Astrophys. J. Supplement*, **48**, 369.

Abbott, D.C., Garmany, C.D., Hansen, C.J., Henrichs, H.F., and Pesnell, W.D. (ed.). 1986, *Pub. Astron. Soc. Pacific*, **98**, 29.

Abbott, D.C., and Hummer, D.G. 1985, *Astrophys. J.*, **294**, 286.

Abbott, D.C., and Lucy, L.B. 1985, *Astrophys. J.*, **288**, 679.

Abbott, D.C., Telesco, C.M., and Wolff, S.C. 1984, *Astrophys. J.*, **279**, 225.

Acker, A. 1984, in *IAU Symp. 105, Observational Test of Stellar Evolution Theory*, ed. A. Maeder and A. Renzini (Dordrecht: Reidel), p. 213.

Allen, C., and Poveda, A. 1968, *Astrophys. J.*, **73**, 586.

Allen, D.A., Wright, A.E., and Goss, W.N. 1976, *Mon. Not. Roy. Astr. Soc.*, **177**, 91.

Anderson, L.S. 1985, *Astrophys. J.*, **298**, 848.

Ando, H. 1985, *Pub. Astron. Soc. Japan*, **37**, 47.

Ando, H. 1986, preprint.

Appenzeller, I. 1970a, *Astron. Astrophys.*, **5**, 355.

Appenzeller, I. 1970b, *Astron. Astrophys.*, **9**, 216.

Appenzeller, I. 1987, in *Instabilities in Luminous Early-Type Stars*, ed. C. de Loore and H.J.G.L.M. Lamers (Dordrecht: Reidel), in press.

Armandroff, T., and Massey, P. 1985, *Astrophys. J.*, **291**, 685.

Auer, L.H., and Mihalas, D. 1972, *Astrophys. J. Supplement*, **24**, 153.

Azzopardi, M., and Breysacher, J. 1979, *Astron. Astrophys.*, **75**, 120.

Azzopardi, M., and Breysacher, J. 1980, *Astron. Astrophys. Supplement*, **39**, 19.

Azzopardi, M., and Vigneau, J. 1975, *Astron. Astrophys. Supplement*, **22**, 285.

Azzopardi, M., and Vigneau, J. 1982, *Astron. Astrophys. Supplement*, **50**, 291.

Baade, D. 1983, *Astron. Astrophys.*, **124**, 283.

Baade, D. 1984a, *Astron. Astrophys.*, **135**, 105.

Baade, D. 1984b, in *Theor. Problems in Stell. Stability and Oscillations, Proc. 25th Liège Int. Astrophys. Colloq.*, ed. A. Noels and M. Gabriel (Liège: Universite de Liège), p. 115.

Baade, D. 1985, *Astron. Astrophys.*, **148**, 59.

Baade, D. 1986, in *Highlights of Astronomy*, Vol. 7, ed. J.-P. Swings (Dordrecht: Reidel), p. 255.

Baade, D., and Ferlet, R. 1984, *Astron. Astrophys.*, **140**, 72.

Balona, L.A., and Crampton, D. 1974, *Mon. Not. Roy. Astr. Soc.*, **166**, 203.

Balona, L.A., and Feast, M.W. 1974, *Mon. Not. Roy. Astr. Soc.*, **167**, 621.

Bappu, M.K.V. 1973, in *IAU Symp. 49, Wolf-Rayet and High Temperature Stars*, ed. M.K.V. Bappu and J. Sahade (Dordrecht: Reidel), p. 59.

Bappu, M.K.V., and Ganesh, K.S. 1968, *Mon. Not. Roy. Astr. Soc.*, **140**, 71.

Bappu, M.K.V., and Sahade, J., (ed.) 1973, *Proc. IAU Symp. 49, Wolf-Rayet and High Temperature Stars*, (Dordrecht: Reidel).

Barlow, M.J. 1982, in *Proc. IAU Symp. 99, Wolf-Rayet Stars: Observations, Physics, Evolution*, ed. C. de Loore and A.J. Willis (Dordrecht: Reidel), p. 149.

Barlow, M.J., Blades, J.C., and Hummer, D.G. 1980, *Astrophys. J. (Letters)*, **241**, L27.

Barlow, M.J., and Cohen, M. 1977, *Astrophys. J.*, **213**, 737.

Barlow, M.J., and Hummer, D.G. 1982, in *Proc. IAU Symp. 99, Wolf-Rayet Stars: Observations, Physics, Evolution*, ed. C. de Loore and A.J. Willis (Dordrecht: Reidel), p. 387.

Barker, P.K., Landstreet, J.D., Marlborough, J.M., and Thompson, I. 1981, *Astrophys. J.*, **250**, 300.

Barker, P.K., and Marlborough, J.M. 1985, *Astrophys. J.*, **288** 329.

Barker, P.K., Marlborough, J.M., and Landstreet, J.D. 1984, in *Proc. of Conf. on Future of Ultraviolet Astronomy Based on Six Years of IUE Research*, ed. J.M. Mead, R.D. Chapman, and Y. Kondo, NASA CP-2349, p. 219.

Barker, T. 1978, *Astrophys. J.*, **219**, 914.

Bates, B., and Halliwell, D. 1986, *Mon. Not. Roy. Astr. Soc.*, **223**, 673.

Beals, C.S., and Plaskett, H.H. 1935, *Trans. IAU 5*, 184.

Bergeron, J. 1977, *Astrophys. J.*, **211**, 62.

Bertelli, G., Bressan, A., and Chiosi, C. 1984, *Astron. Astrophys.*, **130**, 279.

Bertelli, G., and Chiosi, C. 1982, in *Proc. IAU Symp. 99, Wolf-Rayet Stars: Observations, Physics, Evolution*, ed. C. de Loore and A.J. Willis (Dordrecht: Reidel), p. 359.

Bidelman, B. 1977, in *The Impact of Ultraviolet Observations on Spectral Classification*, ed. L. Houziaux, *Highlights of Astronomy*, **4**, Part II, p. 355.

Bisiacchi, G.F., Lopez, J.A., and Firmani, C. 1982, *Astron. Astrophys.*, **107**, 252.

Blaauw, A. 1961, *Bull. Astrn. Inst. Netherlands*, **15**, 265.

Blaauw, A., and Morgan, W.W. 1953, *Bull. Astrn. Inst. Netherlands*, **12**, 76.

Blackwell, D.E., and Shallis, M.J. 1977, *Mon. Not. Roy. Astr. Soc.*, **180**, 177.

Boersma, J. *Bull. Astrn. Inst. Netherlands*, **15**, 291.

Bohannan, B., Abbott, D.C., Voels, S.A., and Hummer, D.G. 1986, *Astrophys. J.*, **308**, 728.

Bohannan, B., Conti, P.S., and Massey, P. 1985, *Astron. J.*, **90**, 600.

Bohannan, B., and Epps, H.W. 1974, *Astron. Astrophys. Supplement*, **18**, 47.

Bohannan, B., and Garmany, C.D. 1978, *Astrophys. J.*, **223**, 908.

Boksenberg, A., Willis, A.J., and Searle, L. 1977, *Mon. Not. Astr. Roy. Soc.*, **180**, 15P.

Bolton, C.T. 1981, in *Proc. IAU Symp. 98, Be Stars*, ed. M. Jaschek and H.-G. Groth (Dordrecht: Reidel), p. 181.

Borsenberger, J., and Stasinska, G. 1982, *Astron. Astrophys.*, **106**, 158.

Bressan, A. Bertelli, G., and Chiosi, C. 1981, *Astron. Astrophys.*, **102**, 25.

Breysacher, J. 1981, *Astron. Astrophys.*, **43**, 203.

Breysacher, J. 1986, *Astron. Astrophys.*, **160**, 185.

Brown, J.C., and Henrichs, H.F., 1987, *Astron. Astrophys.*, **182**, 107.

Brown, P.J.F., Dufton, P.L., Lennon, D.J., Keenan, F.P., and Kilkenny, D. 1986, *Astron. Astrophys.*, **115**, 113.

Bruhweiler, F.C., and Dean, Ch. A. 1983, *Astrophys. J. (Letters)*, **274**, L87.

Bruhweiler, F.C., Kondo, Y., and McCluskey, G. 1981, *Astrophys. J. Supplement*, **46**, 255.

Bruhweiler, F.C., Parsons, S.B., and Wray, J.D. 1982, *Astrophys. J. (Letters)*, **256**, L49.

Brunet, J.P., Imbert, M., Martin, N., Mianes, P., Prevot, L., Rebeirot, E., and Rousseau, J. *Astron. Astrophys. Supplement*, **21**, 109.

Brunish, W., and Truran, J. 1982a, *Astrophys. J.*, **256**, 247.

Brunish, W., and Truran, J. 1982b, *Astrophys. J. Supplement*, **49**, 447.

Burki, G. 1978, *Astron. Astrophys.*, **65**, 357.

Burki, G., Heck, A., Bianchi, L., and Casatella, A. 1982, *Astron. Astrophys.*, **107**, 205.

Burton, W.M., Evans, R.G., and Griffin, W.G. 1975, *Phil. Trans. Roy. Soc. London*, **A279**, 355.

Butler, K., and Simon, K.P. 1985, in *Proc. ESO Workshop, Production and Distribution of C,N,O Elements*, ed. J. Danziger et al., p. 313.

Cardona-Nunes, O. 1978, Ph.D. thesis, University of Colorado.

Carlberg, R.G. 1980, *Astrophys. J.*, **241**, 1131.

Carrasco, L., Aguilar, L.A., and Recillas-Cruz, G. 1982, *Astrophys. J. (Letters)*, **261**, L47.

Carrasco, L., Bisiacchi, G.F., Cruz-Gonzáles, C., Firmani, C., and Costero, R. 1980, *Astron. Astrophys.*, **92**, 253.

Carrasco, L., Costero, R., and Stalio, R. 1981, *Astron. Astrophys.*, **100**, 183.

Carrasco, L., and Creze, M. 1978, *Astron. Astrophys.*, **65**, 279.

Carruthers, G. 1968, *Astrophys. J.*, **151**, 269.

Cassinelli, J.P. 1979, *Ann. Rev. Astron. Astrophys.*, **17**, 275.

Cassinelli, J.P. 1985, in *The Origin of Non-radiative Heating/Momentum in Hot Stars*, ed. A.B. Underhill and A.G. Michalitsianos, NASA CP-2358, p. 2.

Cassinelli, J.P., Castor, J.I., and Lamers, H.J.G.L.M. 1978, *Pub. Astron. Soc. Pacific*, **90**, 496.

Cassinelli, J.P., Hartmann, L., Sanders, W.T., Dupree, A.K., and Myers, R.V. 1983, *Astrophys. J.*, **268**, 205.

Cassinelli, J.P., Mathis, J.S., and Savage, B.D. 1981, *Science*, **212**, 1497.

Cassinelli, J.P., and Olson, G.L. 1979, *Astrophys. J.*, **229**, 304.

Cassinelli, J.P., and Swank, J.H. 1983, *Astrophys. J.*, **271**, 681.

Cassinelli, J.P., Waldron, W.L., Sanders, W.T., Harnden, F.R., Rosner, R., and Vaiana, G.S. 1981, *Astrophys. J.*, **250**, 677.

Castor, J.I. 1986, *Pub. Astron. Soc. Pacific*, **98**, 52.

Castor, J.I. 1987, in *Instabilities in Luminous Early-Type Stars*, ed. C. de Loore and H.J.G.L.M. Lamers (Dordrecht: Reidel), in press.

Castor, J.I., Abbott, D.C., and Klein, R.K. 1975a, *Astrophys. J.*, **195**, 157.

Castor, J.I., and Lamers, H.J.G.L.M. 1979, *Astrophys. J. Supplement*, **39**, 481.

Castor, J.I., McCray, R., and Weaver, R. 1975b, *Astrophys. J. (Letters)*, **200**, L107.

Castor, J.I., and Simon, T. 1983, *Astrophys. J.*, **265**, 304.

Cherepashchuk, A.M. 1973, in *Eclipsing Variable Stars,* ed. V.P. Tsesevich (New York: John Wiley and Sons), p. 243.

Cherepashchuk, A.M. 1981, *Mon. Not. Astr. Roy. Soc.*, **194**, 755.

Cherepashchuk, A.M. 1982, *Astrophys. Space Sci.*, **86**, 299.

Cherepashchuk, A.M., Eaton, J.A., and Khaliullin, Kh. F. 1984, *Astrophys. J.*, **281**, 774.

Chiosi, C., and Maeder, A. 1986, *Ann. Rev. Astron. Astrophys.*, **24**, 329.

Chiosi, C., and Nasi, E. 1974, *Astron. Astrophys.*, **34**, 355.

Chiosi, C., Nasi, E., and Bertelli, G. 1979, *Astron. Astrophys.*, **74**, 62.

Chiosi, C., Nasi, E., and Sreenivasan, S.R. 1978, *Astron. Astrophys.*, **63**, 103.

Chiosi, C., and Stalio, R. 1981, *Proc. IAU Symp. 59, Effects of Mass Loss on Stellar Evolution* (Dordrecht: Reidel).

Chu, Y.-H. 1982, in *Proc. IAU Symp. 99, Wolf-Rayet Stars: Observations, Physics, Evolution,* ed. C.W. de Loore and A.J. Willis (Dordrecht: Reidel), p. 469.

Chu, Y.-H., Cassinelli, J.P., and Wolfire, M.G. 1984, *Astrophys. J.*, **283**, 560.

Chu, Y.-H., and Daod, N.A. 1984, *Pub. Astron. Soc. Pacific*, **96**, 999.

Chu, Y.-H., Treffers, R.R., and Kwitter, K.B. 1983, *Astrophys. J. Supplement*, **53**, 937.

Cloutman, L.D., and Whitaker, R. 1980, *Astrophys. J.*, **237**, 900.

Code, A., Davis, J., Bless, R., and Hanbury-Brown, R. 1976, *Astrophys. J.*, **203**, 417.

Cohen, M., Barlow, M.J., and Kuhi, L.V., 1975, *Astron. Astrophys.*, **40**, 291.

Conti, P.S. 1973a, *Astrophys. J.*, **179**, 161.

Conti, P.S. 1973b, *Astrophys. J.*, **179**, 181.

Conti, P.S. 1975, in *H II Regions and Related Objects*, ed. T.L. Wilson and D. Downes (Heidelberg: Springer-Verlag), p. 207.

Conti, P.S. 1976, *Mem. Soc. Roy. Sci. Liège*, **9**, 193.

Conti, P.S. 1978, *Astron. Astrophys.*, **63**, 225.

Conti, P.S. 1982a, in *Proc. IAU Symp. 99, Wolf-Rayet Stars: Observations, Physics, Evolution*, ed. C. de Loore and A.J. Willis (Dordrecht: Reidel), p. 3.

Conti, P.S. 1982b, in *Proc. IAU Symp. 99, Wolf-Rayet Stars: Observations, Physics, Evolution*, ed. C. de Loore and A.J. Willis (Dordrecht: Reidel), p. 551.

Conti, P.S. 1984, in *Proc. IAU Symp. 105, Observational Tests of the Stellar Theory*, ed. A. Maeder and A. Renzini (Dordrecht: Reidel), p. 233.

Conti, P.S., and Alschuler, W. 1971, *Astrophys. J.*, **170**, 325.

Conti, P.S., and Burnichon, M.L. 1975, *Astron. Astrophys.*, **38**, 467.

Conti, P.S., and Ebbets, D. 1977, *Astrophys. J.*, **213**, 438.

Conti, P.S., and Frost, S.A. 1974, *Astrophys. J. (Letters)*, **190**, L137.

Conti, P.S., and Frost, S.A. 1977, *Astrophys. J.*, **212**, 728.

Conti, P.S., and Garmany, C.D. 1983, *Pub. Astron. Soc. Pacific*, **95**, 411.

Conti, P.S., Garmany, C.D., de Loore, C., and Vanbeveren, D. 1983a, *Astrophys. J.*, **274**, 302.

Conti, P.S., Garmany, C.D., and Hutchings, J.B. 1977a, *Astrophys. J.*, **215**, 561.

Conti, P.S., Garmany, C.D., and Massey, P. 1980, *Astrophys. J.*, **242**, 1063.

Conti, P.S., and Leep, E.M. 1974, *Astrophys. J.*, **193**, 113.

Conti, P.S., Leep, E.M., and Lorre, J. 1977b, *Astrophys. J.*, **214**, 759.

Conti, P.S., Leep, E.M., and Perry, D.N. 1983b, *Astrophys. J.*, **268**, 228.

Conti, P.S., and Massey, P. 1981a, in *Proc. IAU Symp. 59, Effects of Mass Loss on Stellar Evolution*, ed. C. Chiosi and R. Stalio (Dordrecht: Reidel), p. 311.

Conti, P.S., and Massey, P. 1981b, *Astrophys. J.*, **249**, 471.

Conti, P.S., and McCray, R. 1980, *Science*, **208**, 9.

Conti, P.S., and Niemela, V.S. 1976, *Astrophys. J. (Letters)*, **209**, L37.

Conti, P.S., Niemela, V.S., and Walborn, N.R. 1979, *Astrophys. J.*, **228**, 206.

Conti, P.S., Roussel-Dupré, Massey, P., and Rensing, M. 1984, *Astrophys. J.*, **282**, 693.

Cowley, A.P., Crampton, D., Hutchings, J.B., and Thompson, I.B. 1984, *Pub. Astron. Soc. Pacific*, **99**, 968.

Cox, A.N. 1983, in *Astrophysical Processes in Upper Main Sequence Stars*, Swiss Soc. of Astr. and Ap., Geneva Observatory.

Cox, D.P., and Smith, B.W. 1974, *Astrophys. J. (Letters)*, **189**, L105.

Cox, J.P. 1980, *Theory of Stellar Pulsation*, (Princeton: Prnceton Univ. Press).

Crampton, D. 1979, *Astrophys. J.*, **230**, 717.

Crampton, D., and Greasley, J. 1982, *Pub. Astron. Soc. Pacific*, **94**, 31.

Cruz-González, C., Recillas-Cruz, E., Castero, R., Peimbert, M., and Torres-Peimbert, S. 1974, *Revista. Mexicanna de Astron. Y. Astrof.*, **1**, 211.

Danks, A.C., Dennefeld, M., Wamsteker, W., and Shaver, P.A. 1983, *Astron. Astrophys.*, **118**, 301.

Davidson, K., and Kinman, T.D. 1982, *Pub. Astron. Soc. Pacific*, **94**, 634.

Davidson, K., Walborn, N.R., and Gull, T.R. 1982, *Astrophys. J. (Letters)*, **254**, L47.

Davis, R.D., Elliott, K.H., and Measburn, J. 1976, *Mon. Not. Roy. Astr. Soc.*, **81**, 89.

de Jager, C. 1980, *The Brightest Stars*, ed. B.M. McCormack (Dordrecht: Reidel).

de Jager, C., Lamers, H.J.G.L.M., Macchetto, F., and Snow, T.P. 1979, *Astron. Astrophys. (Letters)*, **79**, L28.

de Loore, C. 1980, *Space Sci. Rev.*, **26**, 113.

de Loore, C., De Greve, J.P., and Lamers, H.J.G.L.M. 1977, *Astron. Astrophys.*, **61**, 251.

de Loore, C., De Greve, J.P., and Vanbeveren, D. 1978, *Astron. Astrophys.*, **67**, 373.

de Loore, C., Hellings, D., and Lamers, H.J.G.L.M. 1982, in *Proc. IAU Symp. 99, Wolf-Rayet Stars: Observations, Physics,* *Evolution*, ed. C. de Loore and A.J. Willis (Dordrecht: Reidel), p. 53.

de Loore, C., and Lamers, H.J.G.L.M. (ed.). 1987, *Instabilities in Luminous Early-Type Stars* (Dordrecht: Reidel), in press.

de Loore, C., and Willis, A. (ed.). 1982, *Proc. IAU Symp. 99, Wolf-Rayet Stars: Observations, Physics, Evolution* (Dordrecht: Reidel).

de Loore, C., Willis, A., and Laskarides, D. 1986, in *Proc. IAU Symp. 116, Luminous Stars and Associations in Galaxies* (Dordrecht: Reidel).

Divan, L., Zorec, J., and Andrillat, Y. 1983, *Astron. Astrophys.*, **126**, L8.

Doazan, V. 1986, IAU Circ. 4232.

Doazan, V., Franco, M., Rusconi, L., Sedmak, G., and Stalio, R. 1983, *Astron. Astrophys.*, **128**, 171.

Doazan, V., Grady, C., Kuhi, L.V., Marlborough, J.M., Snow, T.P., and Thomas, R.N. 1982, in *Be Stars*, ed. M. Jaschek and H.-G. Groth (Dordrecht: Reidel), p. 415.

Doazan, V., Grady, C.A., Snow, T.P., Peters, G.J., Marlborough, J.M., Barker, P.K., Bolton, T., Bourdonneau, B., Kuhi, L.V., Lyons, R.W., Polidan, R.S., Stalio, R., and Thomas, R.N. 1985, *Astron. Astrophys.*, **152**, 182.

Doazan, V., Kuhi, L.V., and Thomas, R.N. 1980, *Astrophys. J. (Letters)*, **235**, L17.

Doazan, V., Marlborough, J.M., Morossi, C., Peters, G.J., Rusconi, L., Sedmak, G., Stalio, R., Thomas, R.N., and Willis, A.J. 1986, *Astron. Astrophys.*, **158**, 1.

Doazan, V., Morossi, C., Stalio, R., Thomas, R.N., and Willis, A. 1984, *Astron. Astrophys.*, **131**, 210.

Doazan, V., and Thomas, R.N. 1983.

Doazan, V., and Thomas, R.N. 1986, *IUE Symposium London*, ed. E. Rolfe, ESA SP-263, p. 381.

D'Odorico, S., Baade, D., and Kjär, K. 1981, *Proc. ESO Workshop, The Most Massive Stars* (Munich).

D'Odorico, S., and Benvenuti, P. 1983, *Mon. Not. Roy. Astr. Soc.*, **203**, 157.

D'Odorico, S., Benvenuti, P., and Sabbadin, F. 1978, *Astron. Astrophys.*, **63**, 63.

D'Odorico, S., and Rosa, M. 1981, *Astron. Astrophys.*, **248**, 1015.

D'Odorico, S., and Rosa, M. 1982, *Astron. Astrophys.*, **105**, 410.

Doom, C. 1982a, *Astron. Astrophys.*, **116**, 303.

Doom, C. 1982b, *Astron. Astrophys.*, **116**, 308.

Doom, C., De Greve, J.D., and de Loore, C. 1986, *Astrophys. J.*, **303**, 136.

Drissen, L., Lamontagne, R., Moffat, A.F.J., Bastien, P., and Seguin, M. 1986, *Astrophys. J.*, **304**, 188.

Dufour, R.J. 1984, in *Proc. IAU Symp. 108, Structure and Evolution of the Magellanic Clouds*, ed. S. van den Berg and K.S. de Boer (Dordrecht: Reidel), p. 353.

Durret, F., Bergeron, J., and Boksenberg, A. 1985, *Astron. Astrophys.*, **143**, 347.

Dyck, H.M., Simon T., and Wolstencroft, R.D. 1984, *Astrophys. J.*, **277**, 675.

Ebbets, D.C. 1978, *Astrophys. J.*, **224**, 185.

Ebbets, D.C. 1979a, *Astrophys. J.*, **227**, 510.

Ebbets, D.C. 1979b, *Pub. Astron. Soc. Pacific*, **91**, 804.

Ebbets, D.C. 1980a, *Astrophys. J.*, **235**, 97.

Ebbets, D.C., 1980b, *Astrophys. J.*, **236**, 835.

Ebbets, D.C. 1982, *Astrophys. J. Supplement*, **48**, 399.

Ebbets, D.C., and Conti, P.S. 1977, *Astrophys. J.*, **213**, 438.

Ebbets, D.C., and Conti, P.S. 1982, *Astrophys. J.*, **263**, 108.

Ebbets, D.C., and Savage, B.D. 1982, *Astrophys. J.*, **262**, 234.

Ebbets, D.C., and Wolff, S.C. 1981, *Astrophys. J.*, **243**, 204.

Feast, M.W., and Shuttleworth, M. 1965, *Mon. Not. Roy. Astr. Soc.*, **130**, 245.

Fehrenbach, Ch., Duflot, M., and Acker, A. 1976, *Astron. Astrophys. Supplement*, **24**, 379.

Feitzinger, J.V., Hanuschik, R.W., and Schmidt-Kaler, T. 1983, *Astron. Astrophys.*, **120**, 269.

Fernie, J.D. 1978, *IAU Inf. Bull. Var. Stars*, **1377**.

Firmani, C., Koenigsberger, G., Bisiacchi, G.F., Moffat, A.F.J., and Isserstedt, J. 1980, *Astrophys. J.*, **239**, 607.

FitzGerald, M.D. 1970, *Astron. Astrophys.*, **4**, 234.

Fitzpatrick, E.L. 1982, *Astrophys. J. (Letters)*, **261**, L91.

Fitzpatrick, E.L., Savage, B., and Sitko, M.L. 1982, *Astrophys. J.*, **256**, 578.

Franco, M.L., Kontizas, E., Kontizas, M., and Stalio, R. 1983, *Astron. Astrophys.*, **122**, 9.

Friend, D., and Abbott, D.C. 1986, preprint, *Astrophys. J.*

Friend, D.B., and MacGregor, K.B. 1984, *Astrophys. J.*, **282**, 591.

Gabler, R., Kudritzki, R.P., Pauldrach, A., and Puls, J. 1986, in preparation.

Gamov, G. 1943, *Astrophys. J.*, **98**, 500.

Garmany, C.D. 1982, in *Proc. of Workshop on Observational Basis for Velocity Fields in Stellar Astmospheres*, ed. R. Stalio (Trieste), p. 83.

Garmany, C.D. 1984, *Pub. Astron. Soc. Pacific*, **96**, 779.

Garmany, C.D., and Conti, P.S. 1982, in *Proc. IAU. Symp. 99, Wolf-Rayet Stars: Observations, Physics, Evolution*, ed. C. de Loore and A.J. Willis (Dordrecht: Reidel), p. 105.

Garmany, C.D., and Conti, P.S. 1984, *Astrophys. J.*, **284**, 705.

Garmany, C.D., and Conti, P.S. 1985, *Astrophys. J.*, **293**, 407.

Garmany, C.D., Conti, P.S., and Chiosi, C. 1982, *Astrophys. J.*, **263**, 777.

Garmany, C.D., Conti, P.S., and Massey, P. 1980, *Astrophys. J.*, **242**, 1063.

Garmany, C.D., and Massey, P. 1984, *Bull. Amer. Astron. Soc.*, **16**, 508.

Garmany, C.D., Massey, P., and Conti, P.S. 1984, *Astrophys. J.*, **278**, 233.

Garmany, C.D., Olson, G.L., Conti, P.S., and Van Steenberg, M.E. 1981, *Astrophys. J.*, **250**, 660.

Gathier, R., Lamers, H.J.G.L.M., and Snow, T.P. 1981, *Astrophys. J.*, **247**, 173.

Gehren, T., Husfeld, D., Kudritzki, R.P., Conti, P.S., and Hummer, D.G. 1986, in *Proc. IAU Symp. 116, Luminous Stars and Associations in Galaxies*, ed. C. de Loore, A. Willis, and D. Laskarides (Dordrecht: Reidel).

Giddings, J., Howarth, I.D., and Innes, D. 1986, in preparation.

Gies, D. 1985, Ph.D. thesis, University of Toronto.

Grady, C.A., Bjorkman, K.S., and Snow, T.P. 1987b, preprint.

Grady, C.A., et al. 1982, in *Proc. of Advances in UV Astronomy: Four Years of IUE Research*, ed. Y. Kondo, J.M. Mead, and R.D. Chapman, NASA CP-2238, p. 425.

Grady, C.A., Snow, T.P., and Cash, W.C. 1984, *Astrophys. J.*, **283**, 218.

Grady, C.A., Snow, T.P., and Timothy, J.G. 1983, *Astrophys. J.*, **271**, 691.

Grady, C.A., Sonneborn, G., Wu, Ch.-C., and Henrichs, H.F. 1987a, *Astrophys. J.*, **320**, 376.

Gry, C., Lamers, H.J.G.L.M., and Vidal-Madjar, A. 1984, *Astron. Astrophys.*, **137**, 29.

Güsten, R., and Mezger, D.G. 1983, *Vistas Astron.*, **26**, 159.

Hackwell, J.A., Gehrz, R.D., and Grasdalen, G.L. 1979, *Astrophys. J.*, **234**, 133.

Hackwell, J.A., Gehrz, R.D., Smith, J.R., and Strecker, D.W. 1976, *Astrophys. J.*, **210**, 137.

Haefner, R., Metz, K., and Schoembs, R. 1977, *Astron. Astrophys.*, **55**, 5.

Hamann, W.-R. 1980, *Astron. Astrophys.*, **84**, 342.

Hamann, W.-R. 1981, *Astron. Astrophys.*, **100**, 169.

Hamann, W.-R. 1985a, *Astron. Astrophys.*, **145**, 443.

Hamann, W.-R. 1985b, *Astron. Astrophys.*, **148**, 364.

Hamann, W.-R. 1986, *Astron. Astrophys.*, **160**, 347.

Hamann, W.-R., Gruschinske, J., Kudritzki, R.P., and Simon, K.P. 1981, *Astron. Astrophys.*, **104**, 249.

Hamann, W.-R., and Schmutz, W. 1986, *Astron. Astrophys.*, **166**, L11.

Hammerschlag-Hensberge, G. 1979, IAU Circ. 3391.

Hanbury-Brown, R., Davis, J., and Allen, L.R. 1974, *Mon. Not. Roy. Astr. Soc.*, **167**, 121.

Harnden, F.R., Jr., Branduardi, G., Elvis, M., Gorenstein, P., Grindlay, J., Pye, J.P., Rosner, R., Topka, K., and Vaiana, G.S. 1979, *Astrophys. J. (Letters)*, **234**, L51.

Hayes, D.P. 1975, *Astrophys. J. (Letters)*, **197**, L55.

Hayes, D.P. 1978, *Astrophys. J.*, **219**, 952.

Hayes, D.P. 1984, *Astrophys. J.*, **89**, 1219.

Hayes, D.P. 1986, *Astrophys. J.*, **302**, 403.

Hayes, D.P., and Guinan, E.F. 1984, *Astrophys. J.*, **279**, 721.

Hearn, A. 1975, *Astron. Astrophys.*, **40**, 277.

Heckathorn, J.N., Bruhweiler, F.C., and Gull, T.R. 1982, *Astrophys. J.*, **252**, 230.

Henize, K.G. 1956, *Astrophys. J. Supplement*, **2**, 315.

Henize, K.G., Wray, J.D., and Parsons, S.B. 1981, *Astron. J.*, **86**, 1658.

Henize, K.G., Wray, J.D., Parsons, S.B., and Benedict, G.E. 1975, *Astrophys. J. (Letters)*, **199**, L173.

Henrichs, H.F. 1982, in *Be Stars*, ed. M. Jaschek and H.-G. Groth, (Dordrecht: Reidel), p. 431.

Henrichs, H.F. 1983, Chapter 11 in *Accretion-Driven Stellar X-Ray Sources*, ed. W.H.G. Lewin and E.P.J. van den Heuvel (Cambridge: Cambridge Univ.).

Henrichs, H.F. 1984, in *Proc. 4th European IUE Conf., Rome*, ed. E. Rolfe and B. Battrick, ESA SP-218, p. 43.

Henrichs, H.F. 1986, *Pub. Astron. Soc. Pacific*, **98**, 48.

Henrichs, H.F., Hammerschlag-Hensberge, G., Howarth, I.D., and Barr, P. 1983, *Astrophys. J.*, **268**, 807.

Henrichs, H.F., Hammerschlag-Hensberge, G., and Lamers, H.J.G.L.M. 1980, in *Proc. Second European IUE Conf., Turbingen*, ed. B. Battrick and J. Mort, ESA-SP-157, p. 147.

Henrichs, H.F. and Wakker, B.P. 1987, preprint.

Hidayat, B., Supelli, K.A., and van der Hucht, K.A. 1982, in *Proc. IAU Symp. 99, Wolf-Rayet Stars: Observations, Physics, Evolution*, ed. C. de Loore and A.J. Willis (Dordrecht: Reidel), p. 27.

Hillier, D.J. 1983, Ph.D. thesis, Australian National University, Canberra (microfiche, University of Michigan).

Hillier, D.J. 1984, *Astrophys. J.*, **280**, 744.

Hillier, D.J., Jones, T.J., and Hyland, A.R. 1983, *Astrophys. J.*, **271**, 221.

Hiltner, W.A., and Schild, R.E. 1966, *Astrophys. J.*, **143**, 770.

Hodge, P. 1985, *Pub. Astron. Soc. Pacific*, **97**, 530.

Hodge, P.W., and Wright, F.W. 1967, *The Large Magellanic Cloud* (Washington: Smithsonian Press).

Hodge, P.W., and Wright, F.W. 1977, *The Small Magellanic Cloud* (Washington: Smithsonian Press).

Hollenbock, D., Chu, S.I., and McCray, R. 1976, *Astrophys. J.*, **208**, 458.

Howarth, I.D. 1984, *Mon. Not. Roy. Astr. Soc.*, **206**, 625.

Howarth, I.D., Prinja, R.K., and Willis, A.J. 1984, *Mon. Not. Roy. Astr. Soc.*, **208**, 525.

Hummer, D.G. 1982, *Astrophys. J.*, **257**, 724.

Humphreys, R.M. 1978, *Astrophys. J. Supplement*, **38**, 309.

Humphreys, R.M. 1979, *Astrophys. J.*, **234**, 854.

Humphreys, R.M. 1980a, *Astrophys. J.*, **238**, 65.

Humphreys, R.M. 1980b, *Astrophys. J.*, **241**, 598.

Humphreys, R.M., and Davidson, K. 1979, *Astrophys. J.*, **232**, 409.

Humphreys, R.M., and McElroy, D.B. 1984, *Astrophys. J.*, **284**, 565.

Humphreys, R.M., Nichols, M., and Massey, P. 1985, *Astron. J.*, **90**, 101.

Humphreys, R.M., and Sandage, A. 1980, *Astrophys. J. Supplement*, **44**, 319.

Husfeld, D., Kudritzki, R.-P., Simon, K.P., and Clegg, R.E.S. 1984, *Astron. Astrophys.*, **134**, 139.

Hut, P. 1984a, *Astrophys. J. Supplement*, **55**, 301.

Hut, P. 1984b, in *Proc. IAU Symp. 113, Dynamics of Star Clusters*, ed. J.J. Goodman and P. Hut (Dordrecht: Reidel).

Hutchings, J.B. 1979, *Pub. Astron. Soc. Pacific*, **91**, 361.

Hutchings, J.B. 1982, *Astrophys. J.*, **255**, 70.

Hutchings, J.B., and Cowley, A.P. 1976, *Astrophys. J.*, **206**, 490.

Hutchings, J.B., and Sanyal, A. 1976, *Pub. Astron. Soc. Pacific*, **88**, 279.

Israel, F. 1980, *Astron. Astrophys.*, **90**, 246.

Jeffers, S., Stiff, T., and Weller, W.G., 1985, *Astrophys. J.*, **90**, 1852.

Jenkins, E.B. 1978a, *Astrophys. J.*, **219**, 845.

Jenkins, E.B. 1978b, *Astrophys. J.*, **220**, 107.

Johnson, H.L. 1967, *Astrophys. J. (Letters)*, **150**, L39.

Johnson, H.M. 1978, *Astrophys. J. Supplement*, **36**, 217.

Johnson, H.M., and Hogg, D.E. 1965, *Astrophys. J.*, **142**, 1033.

Joss, P.C., and Rappaport, S. 1984, *Ann. Rev. Astron. Astrophys.*, **22**, 537.

Kahn, F.D. 1981, *Mon. Not. Roy Astr. Soc.*, **196**, 641.

Keel, W.C. 1982, *Pub. Astron. Soc. Pacific*, **94**, 765.

Keenan, F.P., Brown, P.J.F., Dufton, P.L., and Lennon, D.J. 1984, *Astrophys. J. (Letters)*, **279**, L11.

Keenan, F.P., Dufton, P.L., and McKeith, C.D. 1982, *Mon. Not. Roy. Astr. Soc.*, **200**, 673.

Kirbiyik, H., Bertelli, G., and Chiosi, C. 1984, in *Proc. 25th Liège Int. Astrophys. Colloq., Theor. Problems in Stell. Stability and Oscillations,* ed. A. Noels and M. Gabriel (Liège), p. 126.

Kitchin, C.R. 1982, *Early Emission Line Stars* (Bristol: Adam Hilger).

Klare, G., and Neckel, Th. 1977, *Astron. Astrophys. Supplement,* **27**, 215.

Klein, R.I., and Castor, J.I. 1978, *Astrophys. J.*, **220**, 902.

Krolik, J.H., and Raymond, J.C. 1985, *Astrophys. J.*, **298**, 660.

Kudritzki, R.-P. 1976, *Astron. Astrophys.*, **52**, 11.

Kudritzki, R.-P. 1980, *Astron. Astrophys.*, **85**, 174.

Kudritzki, R.-P. 1984, in *Proc. 4th European IUE Conf., Rome*, ed. E. Rolfe and B. Battrick, ESA SP-218, p. 33.

Kudritzki, R.-P. 1985, in *Proc. ESO Workshop, Production and Distribution of C,* *N, O Elements*, ed. J. Danziger et al., p. 277 (invited paper).

Kudritzki, R.-P., and Hummer, D.G. 1986, in *Proc. IAU Symposium 116, Luminous Stars and Associations in Galaxies*, ed. C. de Loore, P. Laskarides, and A.J. Willis (Dordrecht: Reidel), p. 3.

Kudritzki, R.-P., Pauldrach, A., and Puls, J. 1987, *Astron. Astrophys.,* **173**, 293.

Kudritzki, R.-P., Simon, K.P., and Hamann, W.R. 1983, *Astron. Astrophys.,* **118**, 245.

Kuhi, L.V. 1968, *Astrophys. J.,* **152**, 89.

Kunasz, P.B., Hummer, D.G., and Mihalas, D. 1975, *Astrophys. J.,* **202**, 92.

Kunth, D., and Joubert, M. 1985, *Astron. Astrophys.,* **142**, 411.

Kunth, D., and Sargent, W.L.W. 1981, *Astron Astrophys. Letters,* **101**, L5.

Kurucz, R. 1979, *Astrophys. J. Supplement,* **40**, 1.

Kwitter, K.B. 1981, *Astrophys. J.,* **245**, 154.

Kwitter, K.B. 1984, *Astrophys, J.,* **287**, 840.

Kwitter, K.B., and Aller, L.H. 1981, *Mon. Not. Roy. Astr. Soc.,* **195**, 934.

Kwok, S. 1984, in *Effects of Variable Mass Loss on the Local Stellar Environment*, ed. R. Stalio and R.N. Thomas (Trièste: Osservatorio Astronomico), p. 175.

Lacy, C.H. 1977, *Astrophys. J.,* **212**, 132.

Lamers, H.J.G.L.M. 1981, *Astrophys. J.,* **245**, 593.

Lamers, H.J.G.L.M. 1983, in *Diffuse Matter in Galaxies,* ed. J. Audouze, J. Lequeux, M. Levy, and A. Vidal-Madjar (Dordrecht: Reidel), p. 35.

Lamers, H.J.G.L.M. 1987, in *Instabilities in Luminous Early-Type Stars,* ed. C. de Loore and H.J.G.L.M. Lamers (Dordrecht: Reidel), in press.

Lamers, H.J.G.L.M., Gathier, R., and Snow, T.P. 1982, *Astrophys. J.,* **258**, 186.

Lamers, H.J.G.L.M., Korevaar, P., and Cassatella, A. 1985, *Astron Astrophys.,* **149**, 29.

Lamers, H.J.G.L.M., and Morton, D. 1976, *Astrophys. J. Supplement,* **32**, 715.

Lamers, H.J.G.L.M., Waters, L.B., and Wesselius, P.R. 1984, *Astron. Astrophys. (Letters),* **134**, L17.

Lamontagne, R., Moffat, A.E.J., and Seggewiss, W. 1984, *Astrophys. J.,* **277**, 258.

Ledoux, P. 1969, in *La Structure Interne des Etoiles,* XIe Saas Fee Cours, ed. Saas Fee, Geneva Obs.

Leep, E.M., and Conti, P.S. 1979, *Astrophys. J.,* **228**, 224.

Lequeux, J. 1979, *Astron. Astrophys.,* **80**, 35.

Leroy, M., and Lafon, J. 1982, *Astron. Astrophys.,* **106**, 345.

Lesh, J.R. 1968, *Astrophys. J. Supplement,* **151**, 371.

Lesh, J.R. 1982, in *Be Stars With and Without Emission Lines,* ed. A.B. Underhill and V. Doazan, NASA SP-456, p. 147.

Lesh, J.R., and Aizenman, M.L. 1978, *Ann. Rev. Astron. Astrophys.,* **16**, 215.

Lester, J.B. 1973, *Astrophys. J.,* **185**, 253.

Leung, K.-C., Moffat, A.F.J., and Seggewiss, W. 1983, *Astrophys. J.*, **265**, 961.

Leung, K.-C., and Schneider, D.P. 1978a, *Astrophys. J.*, **222**, 924.

Leung, K.-C., and Schneider, D.P. 1978b, *Astrophys. J.*, **224**, 565.

López, J.A., and Walsh, J.R. 1983, *Mon. Not. Roy. Astr. Soc.*, **204**, 129.

Lortet, M.-C, Testor, G., and Niemela, V.S. 1982, in *Proc. IAU Symp. 99, Wolf-Rayet Stars: Observations, Physics, Evolution*, ed. C.W. de Loore and A.J. Willis (Dordrecht: Reidel), p. 473.

Lortet, M.-C., Testor, G., and Niemela, V.S. 1984, *Astron. Astrophys.*, **140**, 24.

Lovy, D., Maeder, A., Noels, A., and Gabriel, M. 1984, *Astron. Astrophys.*, **133**, 307.

Lucke, P.B. 1972, Ph.D. thesis, University of Washington.

Lucke, P.B., and Hodge, P.W. 1970, *Astron. J.*, **75**, 171.

Lucy, L.B. 1976, *Astrophys. J.*, **206**, 499.

Lucy, L.B. 1982a, *Astrophys. J.*, **255**, 278.

Lucy, L.B. 1982b, *Astrophys. J.*, **255**, 286.

Lucy, L.B. 1983, *Astrophys. J.*, **274**, 372.

Lucy, L.B. 1984a, *Astron. Astrophys.*, **140**, 210.

Lucy, L.B. 1984b, *Astrophys. J.*, **284**, 351.

Lucy, L.B. 1986, ESO Preprint 419 to be published in *Proc. IAU Colloq. 89*.

Lucy, L.B., and Solomon, P. 1970, *Astrophys. J.*, **159**, 879.

Lucy, L.B., and White, R.L. 1980, *Astrophys. J.*, **241**, 300.

Lundström, I., and Stenholm, B. 1979, *Astron. Astrophys. Supplement*, **35**, 303.

Lundström, I., and Stenholm, B. 1980, Reports from the Observatory of Lund, Vol. 16.

Lundström, I., and Stenholm, B. 1983, *Inst. Theor. Ap. Oslo Rep.*, No. 59, 7.

Lundström, I., and Stenholm, B. 1984a, *Astron. Astrophys. Supplement*, **56**, 43.

Lundström, I., and Stenholm, B. 1984b, *Astron. Astrophys. Supplement*, **58**, 163.

Lupie, O.L., and Nordsieck, K.H. 1983, *Bull. Amer. Astron. Soc.*, **15**, 876.

Lupie, O.L., and Nordsieck, K.H. 1987, *Astron. J.*, **93**, 214.

MacGregor, K.B., Hartmann, L., and Raymond, J. 1979, *Astrophys. J.*, **231**, 514.

Maeder, A. 1980a, *Astron. Astrophys.*, **92**, 101.

Maeder, A. 1980b, *Astron. Astrophys.*, **90**, 311.

Maeder, A. 1981, *Astron. Astrophys.*, **101**, 385.

Maeder, A. 1982, *Astron. Astrophys.*, **105**, 149.

Maeder, A. 1983, *Astron. Astrophys.*, **120**, 113.

Maeder, A. 1985, *Astron. Astrophys.*, **147**, 300.

Maeder, A. 1986, in *Highlights of Astronomy, Vol. 7*, ed. J.-P. Swings (Dordrecht: Reidel), p. 273.

Maeder, A., and Renzini, A. 1984, *Proc. IAU Symp. 105, Observational Tests of the Stellar Theory* (Dordrecht: Reidel).

Marlborough, J.M., and Peters, G.J. 1982, in *Be Stars*, ed. M. Jaschek and H.-G. Groth (Dordrecht: Reidel), p. 387.

Marlborough, J.M., and Snow, T.P. 1980, *Astrophys. J.*, **235**, 85.

Massa, D., and Savage B.D. 1985, *Astrophys. J.*, **299**, 905.

Massa, D., Savage, B.D., and Fitzpatrick, E.L. 1983, *Astrophys. J.*, **266**, 662.

Massey, P. 1981, *Astrophys. J.*, **246**, 153.

Massey, P. 1984, *Astrophys. J.*, **281**, 789.

Massey P. 1985, *Pub. Astron. Soc. Pacific*, **97**, 5.

Massey, P., Armandroff, T., and Conti, P.S. 1986, in *Proc. IAU Symp. 116, Luminous Stars and Associations in Galaxies*, ed. C. de Loore, A.J. Willis, and P. Laskarides (Dordrecht: Reidel).

Massey, P., and Conti, P.S. 1983a, *Astrophys. J.*, **264**, 126.

Massey, P., and Conti, P.S. 1983b, *Pub. Astron. Soc. Pacific*, **95**, 440.

Massey, P., and Conti, P.S. 1983c, *Astrophys. J.*, **273**, 576.

Massey, P., Conti, P.S., and Niemela, V.S. 1981, *Astrophys. J.*, **246**, 145.

Massey, P., and Hutchings, J.B. 1983, *Astrophys. J.*, **275**, 578.

Massey, P., Lundström, I., and Stenholm, B. 1984, *Pub. Astron. Soc. Pacific*, **96**, 618.

McCluskey, G.E., and Kono, Y. 1981, *Astrophys. J.*, **246**, 464.

McCray, R. 1983, in *Highlights of Astronomy, Vol. 6*, ed. R. West (Dordrecht: Reidel), p. 565.

McCray, R. 1986, in *Spectroscopy of Astrophysical Plasmas*, ed. A. Dalgarno and D. Layzer (Cambridge: Cambridge Univ. Press), in press.

McCray, R., and Snow, T.P., Jr. 1979, *Ann. Rev. Astron. Astrophys.*, **17**, 213.

McKee, C.F., and Ostriker, J.P. 1977, *Astrophys. J.*, **218**, 148.

McKee, C.F., Van Buren, D., and Lazareff, B. 1984, *Astrophys. J. (Letters)*, **278**, L115.

McLean, I.S. 1980, *Astrophys. J. (Letters)*, **236**, L149.

Melnick, J. 1982, in *Proc. IAU Symp. 99, Wolf-Rayet Stars: Observations, Physics, Evolution*, ed. C. de Loore and A.J. Willis (Dordrecht: Reidel), p. 545.

Meylan, G., and Maeder, A. 1982, *Astron. Astrophys.*, **108**, 148.

Mihalas, D. 1972, *Non-LTE Model Atmospheres for B and O Stars*, NCAR TN/STR-76 (Boulder: NCAR).

Mihalas, D. 1984, in *The MK Process and Stellar Classification*, ed. R.F. Garrison (Toronto: David Dunlap Obs.).

Mihalas, D., and Conti, P.S. 1980, *Astrophys. J.*, **235**, 515.

Mihalas, D., and Hummer, D.G. 1973, *Astrophys. J.*, **179**, 827.

Mihalas, D., and Hummer, D.G. 1974, *Astrophys. J. Supplement*, **28**, 343.

Mihalas, D., Hummer, D.G., and Conti, P.S. 1972, *Astrophys. J. (Letters)*, **175**, L99.

Mihalas, D., Kurucz, P.B., and Hummer, D.G. 1975, *Astrophys. J.*, **202**, 465.

Mihalas, D., Kurucz, P.B., and Hummer, D.G. 1976, *Astrophys. J.*, **206**, 515.

Mihalas, D., and Lockwood, G.W. 1972, *Astrophys. J.*, **175**, 757.

Mikkola, S. 1983a, *Mon. Not. Roy. Astr. Soc.*, **203**, 1107.

Mikkola, S. 1983b, *Mon. Not. Roy. Astr. Soc.*, **205**, 733.

Moffat, A.F.J. 1982, in *Proc. IAU Symp. 99, Wolf-Rayet Stars: Observations, Physics, Evolution*, ed. C. de Loore and A.J. Willis (Dordrecht: Reidel), p. 263.

Moffat, A.F.J., Bastien, P., and Drissen, L. 1987a, in *Instabilities in Luminous Early-Type Stars*, ed. C. de Loore and H.J.G.L.M. Lamers (Dordrecht: Reidel), in press.

Moffat, A.F.J., and Michaud, G. 1981, *Astrophys. J.*, **251**, 133.

Moffat, A.F.J., Seggewiss, W., and Shara, M.M. 1985, *Astrophys. J.*, **295**, 109.

Moffat, A.F.J., and Shara, M.M. 1983, *Astrophys. J.*, **273**, 544.

Moffat, A.F.J., Shara, M.M., and Lamontagne, R. 1987b, in *Instabilities in Luminous Early-Type Stars*, ed. C. de Loore and H.J.G.L.M. Lamers (Dordrecht: Reidel), in press.

Morgan, W.W., Whitford, A.E., and Code, A.D. 1950, *Astrophys. J.*, **118**, 318.

Morton, D.C. 1967, *Astrophys. J.*, **147**, 1017.

Morton, D.C. 1969, *Astrophys. J.*, **168**, 629.

Morton, D.C. 1970, *Astrophys. J.*, **160**, 215.

Morton, D.C. 1976, *Astrophys. J.*, **203**, 386.

Mullan, D.J. 1984, *Astrophys. J.*, **283**, 303.

Nandy, K. 1982, in *Ultraviolet Spectral Classification*, ed. A. Heck and B. Battrick ESA SP-182, p. 91.

Niemela, V.S. 1974, Ph.D. thesis, University of Buenos Aires.

Niemela, V.S. 1986, preprint.

Niemela, V.S. 1986, in *Luminous Stars and Associations in Galaxies*, ed. C.W.H. de Loore, A.J. Willis, and P. Laskarides (Dordrecht: Reidel), p. 85.

Niemela, V.S., Mandrini, C.H., and Méndez, R.H. 1986, *Resista. Mex. Astron.*, **11**, 143.

Niemela, V.S., and Moffat, A.F.J. 1982, *Astrophys. J.*, **259**, 213.

Noels, A., Conti, P.S., Gabriel, M., and Vreux, J.-M. 1980, *Astron. Astrophys.*, **92**, 242.

Noels, A., and Gabriel, M. 1981, *Astron. Astrophys.*, **101**, 215.

Noels, A., and Gabriel, M. 1984, in *Proc. 25th Liège Int. Astrophys. Colloq., Theor. Problems in Stell. Stability and Oscillations*, ed. A. Noels and M. Gabriel (Liège), p. 59.

Noels, A., and Scuflaire, R. 1986, *Astron. Astrophys.*, **161**, 125.

Nugis, T. 1975, in *Proc. IAU Symp. 67, Variable Stars and Stellar Electrons*, ed. V.E. Sherwood and L. Plaut (Dordrecht: Reidel), p. 291.

Nussbaumer, H. 1971, *Astrophys. J.*, **170**, 93.

Nussbaumer, H., and Schmutz, W. 1986, *Astron. Astrophys.*, **154**, 100.

Nussbaumer, H., Schmutz, W., Smith, L.J., and Willis, A.J. 1982, *Astron. Astrophys. Supplement*, **47**, 257.

Odegard, N., and Cassinelli, J.P. 1982, *Astrophys. J.*, **256**, 568.

Olson, G.L. 1978, *Astrophys. J.*, **226**, 124.

Olson, G.L. 1982, *Astrophys. J.*, **255**, 267.

Olson, G.L., and Castor, J.I. 1981, *Astrophys. J.*, **244**, 179.

Olson, G.L., and Ebbets, D. 1981, *Astrophys. J.*, **248**, 1021.

Osaki, Y. 1985, in *Seismology of the Sun and the Distant Stars*, ed. D. Gough (Dordrecht: Reidel), p. 453.

Osaki, Y. 1986a, in *Highlights of Astronomy, Vol. 7*, ed. J.-P. Swings (Dordrecht: Reidel), p. 247.

Osaki, Y. 1986b, *Pub. Astron. Soc. Pacific*, **98**, 30.

Osaki, Y. 1987, in *Instabilities in Luminous Early-Type Stars*, ed. C. de Loore and H.J.G.L.M. Lamers (Dordrecht: Reidel), in press.

Osmer, P.S. 1973, *Astrophys. J.* (Letters), **184**, L127.

Osterbrock, D.E., and Cohen, R.D. 1982, *Astrophys. J.*, **261**, 64.

Owocki, S.P., and Rybicki, G.B. 1984, *Astrophys. J.*, **284**, 337.

Owocki, S.P., and Rybicki, G.B. 1985, *Astrophys. J.*, **299**, 265.

Paczynski, B. 1973, in *Proc. IAU Symp. 49, Wolf-Rayet and High Temperature Stars*, ed. M.K.V. Bappu and J. Sahade (Dordrecht: Reidel), p. 143.

Pagel, B.E.J., Edmunds, M.G., Fosbury, R.A.E., and Webster, B.L. 1978, *Mon. Not. Roy. Astr. Soc.*, **184**, 569.

Panagia, N. 1973, *Astron. J.*, **78**, 929.

Panagia, N., and Felli, M. 1975, *Astron. Astrophys.*, **39**, 1.

Panagia, N., and Macchetto, F. 1982a, *Astron. Astrophys.*, **106**, 266.

Panagia, N., and Macchetto, F. 1982b, in *Proc. IAU Colloq. 59, Effects of Mass Loss on Stellar Evolution*, ed. C. Chiosi and R. Stalio (Dordrecht: Reidel), p. 173.

Panek, R.J., and Savage, B.D. 1976, *Astrophys. J.*, **206**, 167.

Parker, R.A.R., Gull, T.R., and Kirchner, R.P. 1979, *An Emission Line Survey of the Milky Way*, NASA SP-434.

Pauldrach, A. 1985, Diplomarbeit, Universitäts-Sternwarte, Universität München.

Pauldrach, A. 1987, *Astron. Astrophys.*, **183**, 295.

Pauldrach, A., Puls, J., Hummer, D.G., and Kudritzki, R.-P. 1985, *Astron. Astrophys. Letters*, **148**, L1.

Pauldrach, A., Puls, J., and Kudritzki, R.P. 1986, *Astron. Astrophys.*, **164**, 86.

Peimbert, M., Torres-Peimbert, S., and Rayo, J. 1978, *Astrophys. J.*, **220**, 516.

Penrod, G.D., 1986, *Pub. Astron. Soc. Pacific*, **98**, 35.

Penrod, G.D., and Vogt, S.S. 1985, *Astrophys. J.*, **299**, 653.

Percy, J.R. 1986, in *Highlights of Astronomy, Vol. 7*, ed. J.-P. Swings (Dordrecht: Reidel), p. 265.

Percy, J.R., and Welch, D.L. 1983, *Pub. Astron. Soc. Pacific*, **95**, 491.

Perry, D.N. 1983, Ph.D. thesis, University of Colorado.

Persi, P., Ferrari-Toniolo, M., and Gradsdalen, G.L. 1983, *Astrophys. J.*, **269**, 625.

Peters, G.J. 1982a, in *Be Stars*, ed. M. Jaschek and H.-G. Groth (Dordrecht: Reidel), p. 353.

Peters, G.J. 1982b, in *Be Stars*, ed. M. Jaschek and H.-G. Groth (Dordrecht: Reidel), p. 401.

Peters, G.J. 1982c, in *Be Stars*, ed. M. Jaschek and H.-G. Groth (Dordrecht: Reidel), p. 411.

Peters, G.J. 1982d, *Astrophys. J. (Letters)*, **253**, L33.

Peterson, D., and Scholz, M. 1970, *Astrophys. J.*, **163**, 51.

Phillips, M.M. 1982, *Mon. Not. Roy. Astr. Soc.*, **198**, 1053.

Pietsch, W., Voges, W., Reppin, C., Trümper, J., Kendziorra, E., and Staubert, R. 1980, *Astrophys. J.*, **237**, 964.

Polidan, R.S., Carone, T.E., and Campbell, C. 1985, *Bull. Amer. Astron. Soc.*, **17**, 554.

Polidan, R.S., and Stalio, R. 1986, in *Proc. IAU Symp. 116, Luminous Stars and Associations in Galaxies*, ed. C. de Loore, A.J. Willis, and P. Laskarides (Dordrecht: Reidel), p. 117.

Pottasch, S.R., Wesselius, P.R., and van Duinen, R.J. 1979, *Astron. Astrophys.*, **77**, 189.

Poveda, A., Ruiz, J., and Allen C. 1967, *Bol. Obs. Tonantzintla y Tacu Baya*, **4**, 860.

Prantzos, N., Doom, C., Arnould, M., and de Loore, C. 1986, *Astrophys. J.*, **304**, 695.

Prévot-Burnichon, M.-L., Prévot, L., Rebeirot, E., Rousseau, J., and Martin, N. 1981, *Astron. Astrophys.*, **103**, 83.

Prinja, R.K. 1984, *Mon. Not. Roy. Astr. Soc.*, **207**, 157.

Prinja, R.K., Henrichs, H.F., Howarth, I.D., and van der Klis, M. 1984, in *Proc. Fourth IUE Conf., Rome*, ed. E. Rolfe and B. Battrick, ESA SP-218, p. 319.

Prinja, R.K., and Howarth, I.D. 1984, *Astron. Astrophys.*, **133**, 110.

Prinja, R.K., and Howarth, I.D. 1985, *Astron. Astrophys.*, **149**, 73.

Prinja, R.K., and Howarth, I.D. 1986, *Astrophys. J. Supplement*, **61**, 357.

Prinja, R.K., Howarth, I.D., and Henrichs, H.F. 1987, *Astrophys. J.*, **317**, 389.

Pylyser, E., Doom, C., and de Loore, C. 1985, *Astron. Astrophys.*, **148**, 379.

Remie, H., and Lamers, H.J.G.L.M. 1982, *Astron. Astrophys.*, **105**, 85.

Roberts, M. 1962, *Astron. J.*, **67**, 79.

Rosa, M., and D'Odorico, S. 1982, *Astron. Astrophys.*, **108**, 339.

Rosendahl, J.D. 1973, *Astrophys. J.*, **186**, 909.

Rousseau, J., Martin, N., Prévot, E., Rebeirot, E., Robin, A., and Brunet, J.P. 1977, *Astron. Astrophys. Supplement*, **31**, 243.

Rufener, F. 1980, *Third Catalogue of Stars Measured in the Geneva Photometric System* (Sauverny, Switzerland: Obs. Genève).

Sadakane, K., Hirata, R., Jugaku, J., Kondo, Y., Matsuoka, M., Tanaka, Y., and Hammerschlag-Hensberge, G. 1985, *Astrophys. J.*, **288**, 284.

Saio, H. 1982, *Astrophys. J.*, **256**, 717.

Sakhibullin, N.A., Auer, L.H., and van der Hucht, K.A. 1982, *Sov. Astron.*, **26**, 563.

Sakhibullin, N.A., and van der Hucht, K.A. 1982, *Astron. Zh.*, **60**, 917 (*Sov. Astron.*, **27**, 529).

Sanduleak, N. 1968, *Astron. J.*, **73**, 246.

Sanduleak, N. 1969, Cerro-Tololo Interamerican Obs. Contrib. No. 89.

Sanduleak, N. 1971, *Astrophys. J. (Letters)*, **164**, L71.

Savage, B.D., Fitzpatrick, E.L., Cassinelli, J.P., and Ebbets, D.C. 1983, *Astrophys. J.*, **273**, 597.

Savage, B.D., and Mathis, J.S. 1979, *Ann. Rev. Astron. Astrophys.*, **17**, 73.

Schild, R.E., Garrison, R.F., and Hiltner, W.A. 1983, *Astrophys. J.*, **88**, 67.

Schneider, D.P., and Leung, K.-C. 1978, *Astrophys. J.*, **223**, 202.

Schönberner, D., Kudritzki, R.-P., and Simon, K.P. 1984, in *Proc. Fourth European IUE Conference*, ESA SP-118, p. 267.

Schwarzchild, M., and Harm, R. 1958, *Astrophys. J.*, **128**, 348.

Scuflaire, R., and Vreux, J.-M. 1987, preprint (submitted to *Astron. Astrophys.*).

Searle, L. 1971, *Astrophys. J.*, **168**, 327.

Seggewiss, W. 1977, in *Proc. IAU Colloq. 42, Interaction of Variable Stars with Their Environment*, ed. R. Kippenhahn, J. Rahe, and W. Strohmeier (Bamberg: Veröff, Remeis-Sternwart), p. 633.

Seggewiss, W., and Moffat, A.F.J. 1979, *Astron. Astrophys.*, **72**, 332.

Severny, A. 1970, *Astrophys. J. (Letters)*, **159**, L73.

Seward, F.D., and Chlebowski, T. 1982, *Astrophys. J.*, **256**, 530.

Shore, S.N., and Sanduleak, N. 1984, *Astrophys. J. Supplement*, **55**, 1.

Shull, J.M. 1980, *Astrophys. J.*, **238**, 860.

Shull, J.M. 1982, in *Regions of Recent Star Formation*, ed. R.S. Roger and P.E. Dewdney (Dordrecht: Reidel), p. 91.

Shull, J.M., and Van Steenburg, M.E. 1985, *Astrophys. J.*, **294**, 599.

Simon, K.P., Jonas, G., Kudritzki, R.-P., and Rahe, J. 1983, *Astron. Astrophys.*, **125**, 34.

Slettebak, A. 1956, *Astrophys. J.*, **124**, 173.

Slettebak, A. 1982, *Astrophys. J. Supplement*, **50**, 55.

Smith, A.M. 1972, *Astrophys. J.*, **176**, 405.

Smith, L.F. 1967, *Astrophys. J.*, **72**, 829.

Smith, L.F. 1968a, *Mon. Not. Roy. Astr. Soc.*, **140**, 409.

Smith, L.F. 1968b, *Mon. Not. Roy. Astr. Soc.*, **141**, 317.

Smith, L.F. 1973, in *Proc. IAU Symp. 49, Wolf-Rayet and High Temperature Stars*, ed. M.K.V. Bappu and J. Sahade (Dordrecht: Reidel), p. 15.

Smith, L.F., and Aller, L.H. 1969, *Astrophys. J.*, **157**, 1245.

Smith, L.F., and Aller, L.H. 1971, *Astrophys. J.*, **164**, 275.

Smith, L.F., and Kuhi, L.V. 1970, *Astrophys. J.*, **162**, 535.

Smith, L.F., and Kuhi, L.V. 1981, JILA Publication.

Smith, L.J., Lloyd, C., and Walker, E.N. 1985, *Astron. Astrophys.*, **146**, 307.

Smith, L.J., and Willis, A.J. 1982, *Mon. Not. Roy. Astr. Soc.*, **201**, 451.

Smith, M.A. 1978, *Astrophys. J.*, **224**, 927.

Smith, M.A. 1981, *Astrophys. J.*, **248**, 214.

Smith, M.A. 1986a, *Pub. Astron. Soc. Pacific*, **98**, 33.

Smith, M.A. 1986b, *Astrophys. J.*, **304**, 728.

Smith, M.A., and Ebbets, D. 1981, *Astrophys. J.*, **247**, 158.

Smith, M.A., and Penrod, G.D. 1985, in *Relations Between Chromospheric-Coronal Heating and Mass Loss in Stars*, ed. R. Stalio and J. Zirker (Trièste), p. 394.

Snow, T.P. 1977, *Astrophys. J.*, **217**, 760.

Snow, T.P. 1979, in *Mass Loss and Evolution of O-Type Stars*, ed. P.S. Conti and C.W.H. de Loore (Dordrecht: Reidel), p. 65.

Snow, T.P., Cash, W., and Grady, C.A. 1981, *Astrophys. J. (Letters)*, **244**, L19.

Snow, T.P., and Hayes, D.P. 1978, *Astrophys. J.*, **226**, 897.

Snow, T.P., and Morton, D.C. 1976, *Astrophys. J. Supplement*, **32**, 429.

Snow, T.P., Oegerle, W.R., and Polidan, R.S. 1980b, *Astrophys. J.*, **242**, 1077.

Snow, T.P., Wegner, G.A., and Kunasz, P.B. 1980a, *Astrophys. J.*, **238**, 643.

Sobolev, V.V. 1960, *The Moving Envelopes of Stars* (Cambridge: Harvard Univ. Press).

Sonneborn, G., Grady, C., Wu, Ch.-C., Hayes, D., Guinan, E., Barker, P., and Henrichs, H.F., 1987, *Astrophys. J.*, in press.

Stahl, O. 1986, ESO Scient. Preprint, No. 428.

Stahl, O., Wolf, B., Klare, G., Cassatelli, A., Krauter, J., Persi, D., and Ferrari-Toniolo, M. 1983, *Astron. Astrophys.*, **127**, 49.

Stalio, R., Sedmak, G., and Rusconi, L. 1981, *Astron. Astrophys.*, **101**, 168.

Stalio, R., and Zirker, J.B. (ed.), 1985, *Relations Between Chromospheric-Coronal Heating and Mass Loss in Stars*, 1984 Workshop, Sacramento Peak Observatory, Sunspot, New Mexico.

Stasinska, G. 1982, *Astron. Astrophys. Supplement*, **48**, 299.

Steigman, G., Strittmatter, P.A., and Williams, R.E. 1975, *Astrophys. J.*, **198**, 575.

Stickland, D.J., Bromage, G.E., Budding, E., Burton, W.M., Howarth, I.D., Jameson, R., Sherrington, M.R., and Willis, A.J. 1984, *Astron. Astrophys.*, **134**, 45.

Stone, R.C. 1978, *Astron. J.*, **83**, 393.

Stone, R.C. 1979, *Astrophys. J.*, **232**, 520.

Stone, R.C. 1981, *Astron. J.*, **86**, 544.

Stone, R.C. 1982, *Astron. J.*, **87**, 90.

Stothers, R., and Chin, C.W. 1977, *Astrophys. J.*, **211**, 189.

Stothers, R., and Chin, C.W. 1980, *Astrophys. J.*, **240**, 885.

Takada, M. 1977, *Pub. Astron. Soc. Japan*, **29**, 439.

Thevenin, F., and Pitault, A. 1982, *Astron. Astrophys.*, **108**, 195.

Thomas, R.N. 1973, *Astron. Astrophys.*, **29**, 297.

Thomas, R.N. 1982, *Astrophys. J.*, **263**, 870.

Thomas, R.N. 1983, *Stellar Atmospheric Structural Patterns*, NASA SP-471.

Thomas, R.N., and Stalio, R. 1984, *Effects of Variable Mass Loss on the Local Stellar Environment* (Trièste: Osservatorio Astronomico).

Tobin, W. 1983, *Astron. Astrophys.*, **125**, 168.

Tobin, W., and Kilkenny, D. 1981, *Mon. Not. Roy. Astr. Soc.*, **194**, 937.

Torres, A.V. 1985, Ph.D. thesis, University of Colorado.

Torres, A.V., and Conti, P.S. 1984, *Astrophys. J.*, **280**, 181.

Torres, A.V., Conti, P.S., and Massey, P. 1986, *Astrophys. J.*, **379**, 309.

Turner, D.G. 1982, in *Proc. IAU Symp. 99, Wolf-Rayet Stars: Observations, Physics, Evolution*, ed. C. de Loore and A.J. Willis (Dordrecht: Reidel), p. 57.

Underhill, A.B. 1968, *Ann. Rev. Astron. Astrophys.*, **6**, 39.

Underhill, A.B. 1975, *Astrophys. J.*, **199**, 691.

Underhill, A.B. 1982, *Astrophys. J.*, **263**, 741.

Underhill, A.B. 1984, *Astrophys. J.*, **285**, 668.

Underhill, A.B. 1985, *Astron. Astrophys.*, **148**, 431.

Underhill, A.B., and Doazan, V. 1982, *B Stars With and Without Emission Lines*, NASA SP-456.

Underhill, A.B., and Fahey, R.P. 1984, *Astrophys. J.*, **280**, 712.

Underhill, A.B., and Michalitsianos, A.G. 1985, NASA CP-2358.

Unno, W., Osaki, Y., Ando, H., and Shibahashi, H. 1979, *Non-radial Oscillations of Stars* (Tokyo: Univ. Tokyo Press).

van Albada, T.S. 1968, *Bull. Astron. Inst. Netherlands*, **20**, 57.

Vanbeveren, D., and Conti, P.S. 1980, *Astron. Astrophys.*, **88**, 230.

Vanbeveren, D., van Rensbergen, W., and de Loore, C. 1982, in *Proc. ESO Workshop, The Most Massive Stars*, ed. S. D'Odorico, D. Baade, and K. Kjär (ESO Publication), p. 199.

Van Buren, D. 1985, *Astrophys. J.*, **294**, 567.

van den Berg, S., and de Boer, K.S. 1984, *Proc. IAU Symp. 108, Structure and Evolution of the Magellanic Clouds* (Dordrecht: Reidel).

van den Heuvel, E.P.J. 1978, in *Physics and Astrophysics of Neutron Stars and Black Holes*, ed. R. Giacconi and R. Ruffini (Amsterdam: North-Holland), p. 828.

van den Heuvel, E.P.J. 1985, in *Birth and Evolution of Massive Stars and Stellar Groups,* ed. W. Boland and H. Van Woerden (Dordrecht: Reidel), p. 107.

van der Hucht, K.A., Conti, P.S., Lundström, I., and Stenholm, B. 1981, *Space Sci. Rev.,* **28,** 227.

van der Hucht, K.A., Jurriens, T.A., Olnon, F.M., Thé, P.S., Wesselius, P.R., and Williams, P.M. 1985, *Astron. Astrophys. Letters,* **145,** L13.

van der Klis, M., and Bonnet-Bidaud, J.M. 1984, *Astron. Astrophys.,* **135,** 155.

van Genderen, A.M. 1985, *Astron. Astrophys.,* **151,** 349.

van Genderen, A.M. 1986, *Astron. Astrophys.,* **157,** 163.

van Genderen, A.M., Alphenaar, P., van der Bij, M.D.P., Deul, E.R., van Driel, W., van Heerde, G.M., de Lange, L., van Leeuwen, F., Meys, J.J.M., Oppe, J., Thé, P.S., and Wiertz, M.J.J. 1985a, *Astron. Astrophys. Supplement,* **61,** 213.

van Genderen, A.M., van den Boogart, A.K., Brand, J., de Bruyn, J.W., van der Bij, M.D.P., Burger, J.H., Deul, E.R., van Heerde, G.M., Hooimeyer, J.R.A., Kraakman, H., de Lange, L., Lemstra, G., Oppe, J., and Wiertz, M.J.J. 1985b, *Astron. Astrophys. Supplement,* **62,** 291.

Viallefond, F. 1984a, in *Proc. Workshop at RAL, Spectral Evolution of Galaxies,* RAL-84-008, ed. P.M. Gondhal-Ekar, p. 28.

Viallefond, F. 1984b, in *Wolf-Rayet Stars: Progenitors of Supernovae?,* ed M.-C. Lortet and A.P. Tault (a workshop held at Observatoire de Paris, Meudon), p. V23.

Vogt, S.S., and Penrod, G.D. 1983a, *Astrophys J.,* **280,** 712.

Vogt, S.S., and Penrod, G.D. 1983b, *Astrophys. J.,* **275,** 661.

Vreux, J.-M. 1985a, *Astron. Astrophys.,* **143,** 209.

Vreux, J.-M. 1985b, *Pub. Astron. Soc. Pacific,* **97,** 274.

Vreux, J.-M. 1987, in *Instabilities in Luminous Early-Type Stars,* ed. C. de Loore and H.J.G.L.M. Lamers (Dordrecht: Reidel), in press.

Vreux, J.-M., and Andrillat, Y. 1979, *Astron. Astrophys.,* **75,** 93.

Vruex, J.-M., and Andrillat, Y., and Gosset, E. 1985, *Astron. Astrophys.,* **149,** 337.

Vreux, J.-M., and Conti, P.S. 1979, *Astrophys. J.,* **228,** 220.

Vreux, J.-M., Dennefeld, M., and Andrillat, Y. 1982, *Astron. Astrophys. Letters,* **113,** L10.

Vreux, J.-M., Dennefeld, M., and Andrillat, Y. 1983, *Astron. Astrophys. Supplement,* **54,** 437.

Wakker, B.P., and Henrichs, H.F. 1987, in preparation.

Walborn, N.R. 1970, *Astrophys. J. (Letters),* **161,** L149.

Walborn, N.R. 1971a, *Astrophys. J. (Letters),* **164,** L67.

Walborn, N.R. 1971b, *Astrophys. J. (Letters),* **167,** L31.

Walborn, N.R. 1971c, *Astrophys. J. Supplement,* **23,** 257.

Walborn, N.R. 1972, *Astron. J.*, **77**, 312.

Walborn, N.R. 1973a, *Astrophys. J. (Letters)*, **182**, L21.

Walborn, N.R. 1973b, *Astron. J.*, **78**, 1907.

Walborn, N.R. 1974, *Astrophys. J.*, **189**, 269.

Walborn, N.R. 1976, *Astrophys. J.*, **205**, 419.

Walborn, N.R. 1977, *Astrophys. J.*, **215**, 53.

Walborn, N.R. 1979, in *Proc. IAU Colloq. 47, Spectral Classification of the Future, Ricerche Astron.*, **9**, 337.

Walborn, N.R. 1980, *Astrophys. J. Supplement*, **44**, 535.

Walborn, N.R. 1982a, *Astrophys. J.*, **256**, 452.

Walborn, N.R. 1982b, *Astrophys. J. (Letters)*, **254**, L15.

Walborn, N.R. 1982c, *Astron. J.*, **87**, 1300.

Walborn, N.R. 1983a, *Astrophys. J.*, **265**, 716.

Walborn, N.R. 1983b, *Astrophys. J. (Letters)*, **267**, L59.

Walborn, N.R. 1984a, in *Proc. IAU Symp. 108, Structure and Evolution of the Magellanic Clouds*, ed. S. van den Berg and K.S. de Boer (Dordrecht: Reidel), p. 239.

Walborn, N.R. 1984b, in *Proc. IAU Symp. 108, Structure and Evolution of the Magellanic Clouds*, ed. S. van den Berg and K.S. de Boer (Dordrecht: Reidel), p. 243.

Walborn, N.R. 1984c, in *The MK Process and Stellar Classification*, ed. R.F. Garrison (Univ. Toronto: David Dunlap Obs.), p. 346.

Walborn, N.R. 1986, in *Luminous Stars and Associations in Galaxies*, ed. C. de Loore, P. Laskarides, and A.J. Willis (Dordrecht: Reidel), p. 185.

Walborn, N.R., Conti, P.S., and Vreux, J.-M. 1980, *Pub. Astron. Soc. Pacific*, **92**, 284.

Walborn, N.R., Nichols-Bohlin, J., and Panek, R.J. 1985, *IUE Atlas of O-Type Spectra from 1200 to 1900 Å*, NASA RP-1155.

Walborn, N.R., and Panek, R.J. 1984a, *Astrophys. J. (Letters)*, **280**, L27.

Walborn, N.R., and Panek, R.J. 1984b, *Astrophys. J.*, **286**, 718.

Walborn, N.R., and Panek, R.J. 1985, *Astrophys. J.*, **291**, 806.

Waldron, W.L. 1984, *Astrophys. J.*, **282**, 256.

Wampler, E.J. 1982, *Astron. Astrophys.*, **114**, 165.

Weaver, R., McCray, R., Castor, J.I., Shapiro, P., and Moore, R. 1977, *Astrophys. J.*, **218**, 377.

Weber, S.V. 1981, *Astrophys. J.*, **243**, 954.

Wegner, G.A., and Snow, T.P. 1978, *Astrophys. J. (Letters)*, **226**, L25.

West, D.K. 1972, in *Scientific Results from OAO-2*, ed. A.D. Code, NASA SP-310, p. 441.

Westerlund, B.E., Azzopardi, M., and Breysacher, J. 1986, *Astron. Astrophys. Supplement*, **65**, 79.

Westerlund, B.E., Azzopardi, M., Breysacher, J., and Lequeux, J. 1983, *Astron. Astrophys.*, **123**, 159.

White, N.E., Kallman, T., and Swank, J.H. 1983, *Astrophys. J.*, **269**, 264.

White, R.L. 1985, *Astrophys. J.*, **289**, 698.

Williams, P.M., Adams, D.J., Arakaki, S., Beattie, D.H., Born, J., Lee, T.J., Robertson, D.J., and Stewart, J.M. 1980, *Mon. Not. Roy. Astr. Soc.*, **192**, 25.

Williams, P.M., Beattie, D.H., Lee, T.J., Stewart, J.M., and Antonopoulou, E. 1978, *Mon. Not. Roy. Astr. Soc.*, **185**, 467.

Williams, P.M., Longmore, A.J., van der Hucht, K.A., Talevera, A., Wamsteker, W.M., Abbott, D.C., and Telesco, C.M. 1985, *Mon. Not. Roy. Astr. Soc.*, **215**, 21.

Williams, P.M., van der Hucht, K.A., Wamsteker, W.M., Gaballe, T.R., Garmany, C.D., and Pollock, A.M.T. 1987, in *Instabilities in Luminous Early-Type Stars*, ed. C. de Loore and H.J.G.L.M. Lamers (Dordrecht: Reidel), in press.

Willis, A.J. 1981, in *Proc. IAU Symp. 59, Effects of Mass Loss on Stellar Evolution*, ed. C. Chiosi and R. Stalio (Dordrecht: Reidel), p. 27.

Willis, A.J. 1982a, in *Proc. IAU Symp. 99, Wolf-Rayet Stars: Observations, Physics, Evolution*, ed. C. de Loore and A.J. Willis (Dordrecht: Reidel), p. 87.

Willis, A.J. 1982b, *Mon. Not. Roy. Astr. Soc.*, **198**, 897.

Willis, A.J., Conti, P.S., Garmany, C.D., and Howarth, I.D. 1986, in *Proc. IAU Symp. 116, Luminous Stars and Associations in Galaxies*, ed. C. de Loore, A.J. Willis, and P. Laskarides (Dordrecht: Reidel), p. 259.

Willis, A.J., and Stickland, D.J. 1980, *Mon. Not. Roy. Astr. Soc.*, **190**, 27.

Willis, A.J., van der Hucht, K.A., Conti, P.S., and Garmany, C.D. 1986, *Astron. Astrophys. Supplement*, **63**, 417.

Willis, A.J., and Wilson, R. 1976, *Astron. Astrophys.*, **47**, 429.

Willis, A.J., and Wilson, R. 1978, *Mon. Not. Roy. Astr. Soc.*, **182**, 559.

Willson, L.A. 1986, *Pub. Astron. Soc. Pacific*, **98**, 35.

Wolf, B. 1986, in *Proc. IAU Symp. 116, Luminous Stars and Associations in Galaxies*, ed. C. de Loore, A.J. Willis, and P. Laskarides (Dordrecht: Reidel), p. 151.

Wolf, B., and Stahl, O. 1983, *ESO Messenger*, No. 33, p. 11.

Wolff, R.J. 1963, *Pub. Astron. Soc. Pacific*, **75**, 485.

Wray, J.D., and Corso, G.J. 1972, *Astrophys. J.*, **172**, 577.

Wright, A.E., and Barlow, M.J. 1975, *Mon. Not. Roy. Astr. Soc.*, **170**, 41.

York, D.G., Vidal-Madjar, A., Laurent, C., and Bonnet, R. 1977, *Astrophys. J. (Letters)*, **213**, L61.

PART III

Another Perspective on O, Of, and Wolf-Rayet Stars, Emphasizing Model Atmospheres and Possibilities for Atmospheric Heating

Written by

Anne B. Underhill

6

UNDERSTANDING THE O AND WOLF-RAYET STARS

I. INTRODUCTION

Observations of stellar spectra may be transformed into information about the physical state of the star only when they are used with theory, however simple or complex. In Part III, we shall first attempt to understand the link between observed spectra and theory which presently exists. After that we shall explore some ways in which the theory may be improved to provide a better understanding of the full set of observations which have been made. A complete understanding of the implications of stellar spectra requires not only linking the theory of stellar spectra to the observations of stellar spectra, but also the development of a logically firm link to the theory of stellar evolution.

The tasks which we face are the following. From the energy distribution of an O or Wolf-Rayet star, observed over a wide range of wavelength and, at least in part, at high spectral resolution: (1) deduce the composition of and physical conditions in the layers of the star which send radiation to us, and (2) deduce the gross properties of the star, including the mass, radius, internal structure, and stage of evolution. We wish to consider the question: Does a spectral type based on the relative strengths of a few empirically selected lines in the visible range uniquely determine the gross properties of the star, including its stage of evolution? A final task is to demonstrate what kind of nonradiative phenomena occur in the atmospheres of O and Wolf-Rayet stars and to relate the occurrence of these phenomena to the stage of evolution of the star.

A. Method of Analysis

The traditional method of finding answers to the tasks posed above is to solve each part of the problem in parallel. One develops models for the atmosphere of the star, based on appropriate parts of the physics of gas and radiation, and predicts stellar spectra which are then compared with observed stellar spectra. From the degree of agreement between the observed and predicted spectra, one deduces the values of the parameters which define a representative model atmosphere. The classical parameters are the composition of the atmosphere, the effective temperature, T_{eff}, and log g, where g is the value of the acceleration of gravity in the boundary layers of the star. In classical model atmospheres, the ensemble of gas and radiation is assumed to be in radiative and hydrostatic equilibrium.

To understand how a real star may have evolved, one develops a model for the star by assuming a mass and a composition and applying knowledge obtained from physics to deduce how energy is generated in the center of the star, how much energy escapes from the boundary layer of the model star as radiation (which determines the effective temperature of the star), and what the radius of the star is. This information is determined at a number of stages in the lifetime of the model star. Finally, one compares the values for the effective temperature, radius, and mass of a real star determined from analysis of the spectrum of the star and identifies each real star with a model star at some particular stage of evolution.

However, we shall see that important parts of the spectra of O and Wolf-Rayet stars cannot be understood by means of the traditional methods of analysis in which only classical concepts are used for defining model atmospheres and model stars. An important goal of Part III will be to identify what cannot be understood at this time. We shall try to deduce what types of nonradiative phenomena are occurring, and to assess the significance of what is observed for understanding the evolution of a massive star.

B. Anatomy of a Star

From the earliest studies of the physics of stars, model stars have been divided into two parts: (1) the interior, where energy is generated and transmitted outward, and (2) the atmosphere, which forms the boundary layer of the star and from which radiation is emitted into the surrounding interstellar medium. When it was realized that the energy radiated by a star during most of its lifetime is generated by means of nuclear reactions, it became convenient to divide the interior of a star into two parts: (1) the core, which is the place where energy is generated, and (2) the envelope, which is the rest of the interior of the star. The principal function of the envelope of a star is to transmit energy to the boundary layer or atmosphere of the star.

The traditional, or classical, model atmosphere is defined in terms of plane parallel layers of gas which are in radiative and hydrostatic equilibrium. The amount of energy in the form of radiation furnished to each square cm of surface of the atmosphere from the envelope of the star is measured by the parameter effective temperature; the pressure (density) structure of the atmosphere is found by constraining the gas to be in hydrostatic equilibrium under the inward acceleration due to the attraction of gravity arising from the mass of the star and the outward acceleration due to radiation pressure. The parameter determining the pressure structure is log g, where $g = GM_*/R_*^2$. The distribution of atoms and ions over their possible energy states may be calculated either by using the equations valid for thermodynamic equilibrium at the local temperature and density (LTE) or by solving the equations of statistical equilibrium (non-LTE).

We shall see that some parts of the spectra of O and Wolf-Rayet stars seem to be formed in parts of the atmosphere where additional energy that is not from the radiation stream is deposited. This source of energy is nonradiative. That is, it is not included in the energy stream measured by means of the parameter, T_{eff}. To model these parts of the atmosphere, one has to start from different precepts than those which are valid for classical model atmospheres. In Chapter 8, we shall discuss some useful ways to proceed.

Simple considerations indicate that the parts of the atmosphere in which the effects of the deposition of nonradiative energy and momentum are important lie outside most of the part which can be represented acceptably by means of classical model atmospheres. Consequently, for clarity, it is useful to divide the atmosphere of a star into two parts: the inner part, or *photosphere*, which may be represented adequately by classical model atmospheres, and the outer atmosphere, or *mantle*. In the mantle, the deposition of nonradiative energy and momentum is an important factor which controls the physical state of this part of the atmosphere.

The word "mantle" has been chosen (see, for instance, Underhill, 1980b, 1981, and Underhill, 1980a, where the word "sheath" was used in place of mantle) as a name for this part of the atmosphere, whose special properties have only recently been recognized for the hot stars. This choice has been made because a major meaning of the word "mantle" is something which covers, envelopes, or conceals. We shall see that, in the case of the O and Wolf-Rayet stars, spectral features formed in the mantle frequently cover and conceal features formed in the photosphere. The mantle of a star envelopes the rest of the star, including the photosphere. A further discussion of this point may be found in Underhill and Doazan (1982), page 8.

C. Spectroscopic Criteria

Spectroscopic features suitable for giving information about the physical state of the photosphere, or of the mantle, must be features which are formed predominantly in the photosphere, or in the mantle, respectively. The Eddington-Barbier principle states that the light seen in a spectroscopic feature comes predominantly from that layer of the atmosphere in which the monochromatic optical depth for wavelengths typical of the feature under study is about 0.67. This statement is valid for stationary atmospheres. It is a useful diagnostic tool for obtaining a first idea about which parts of a model atmosphere are significant for determining the emergent flux in selected absorption lines and in the continuum.

In the case of model atmospheres having T_{eff} greater than about 25000 K and log g of the order of 4.0, it is easy to show that the opacity in the prominent lines of the visible spectral region used for spectral classification is such that the geometric level at which the monochromatic optical depth in the wavelengths of these lines is about 0.67 lies well outside the geometric level at which the monochromatic optical depth in the continuum between 4000 and 6000 Å reaches about 0.67. This means that the shape of the continuous spectrum in the visible range gives information about the physical state of the model in a rather different part of the model than the part which is seen by means of the prominent absorption lines.

It is chiefly the prominent absorption and emission features between 4000 and 6000 Å which have been chosen empirically to define the spectral types of O and Wolf-Rayet stars. A little work has been done to define the spectral types of these stars by means of the shape of the continuous spectrum in one or more spectral regions. Because of the behavior described in the preceding paragraph, O spectral types based on absorption lines correspond to the physical state of a different part of the atmosphere than do O types determined from the shape of the continuum in the visible spectral range. Wolf-Rayet spectral types are determined only from the relative strengths of emission lines. These spectral types, therefore, correspond to the physical conditions only in the region of the atmosphere where emission lines are formed.

If the traditional concepts for making model atmospheres are valid for all parts of the atmosphere contributing to the visible spectrum, then the interpretation of spectral types assigned on the basis of the relative strengths of the absorption lines should yield similar information to that found from an interpretation of the shape of the continuous spectrum and from the total amount of energy emerging from the star as radiation. In practice, it is found that the interpretations of the strengths of the absorption lines of O stars do not always lead to the same effective temperatures as are found from interpretations of the shape of the continuous spectrum or from the total amount of energy radiated by the star.

When one considers stellar spectra which contain emission features, such as the spectra of Of and Wolf-Rayet stars, the models valid for interpreting only absorption lines and the continuum must be changed. This is because traditional classical model atmospheres do not predict the occurrence of significant emission lines. One finds that the emission features are formed predominantly in a part of the atmosphere which is different from that which produces the continuum and the normal absorption lines. This additional part may be of wide extent and/or have a high temperature.

These facts are the reason why it is most desirable to begin a discussion of the meaning of the spectra of O and Wolf-Rayet stars by postulating the presence of a two-part atmosphere.

Here we postulate that O and Wolf-Rayet stars possess a photosphere which can be represented acceptably by traditional, classical model atmospheres, and that outside the photosphere there is a mantle which is easily seen by means of intrinsically strong spectral lines from abundant atoms and ions. In the mantle, the physical state is determined chiefly by the consequences of the deposition of nonradiative energy and momentum.

Naturally, the atmosphere of a real star cannot be unequivocally divided into two distinct, noninteracting parts. However, it must be stressed that one will obtain a clearer idea of the physical state in the photosphere of the star if one restricts oneself to the analysis of spectroscopic features which can be shown to be formed chiefly in the photosphere than if one analyzes a random selection of spectroscopic features. Similarly, one will obtain a clearer idea of the properties of the mantle if one interprets only spectral features formed predominantly in the mantle in order to determine the physical state of the mantle.

Weak spectral lines from ions of low abundance and the continuous spectrum are good criteria to analyze in order to obtain information about the photosphere. Strong lines from abundant ions will yield information about the physical state of the mantle. Analysis of emission lines also will yield information about the mantle of a hot star. In the rest of Part III, an attempt is made to be specific about which part of the atmosphere is being studied when interpretation is made of selected spectroscopic features. As a first approximation, it is assumed that analysis of the continuous spectrum will lead to information about the photosphere, and that analysis of the equivalent widths of lines from abundant ions most likely gives information about conditions in the mantle.

D. Analogy to the Sun

In order to direct our thoughts, we shall use the Sun as a guide to what may occur in the atmospheres of stars. The atmosphere of the Sun, viewed globally, may be considered to have a two-part structure: (1) a photosphere which may be modeled rather successfully by classical methods, and (2) a mantle which lies outside the photosphere. The mantle is not easily observed from the surface of the Earth unless the disk of the Sun is occulted or one observes in the nonvisual spectrum. The mantle of the Sun has many properties like those of the mantles of O and Wolf-Rayet stars; it may serve as a model for stellar mantles.

Observation of the Sun has shown that all parts of the solar mantle have fine structure. Conditions are neither constant nor uniform throughout the mantle of the Sun. Many changing structures such as spicules, prominences, coronal arches, and coronal holes are recognized. Although a fully satisfactory theory of the formation, heating, and change of the full solar mantle does not exist, it seems clear that the presence of weak magnetic fields in low-density ionized gas and differential motions in the photosphere are important factors for causing the observed changes of the physical state of the various parts of the solar mantle. This point of view has been emphasized by Rosner et al. (1978a, 1978b), and it is frequently considered when attempts are made to explain what is seen for the Sun (see, for instance, several of the chapters in Jordan, 1981).

In Chapters 8 and 9, we explore the working hypothesis that weak magnetic fields and differential motion in the photosphere work together to provide at least some of the spectroscopic phenomena seen for O and Wolf-Rayet stars.

Until the properties of the mantles of O and Wolf-Rayet stars are sufficiently well understood for us to know exactly what temperatures occur in the mantles, what the density pattern is, and how flow originates and is accelerated, it seems premature to use words such as "chromosphere" and "corona" for referring to the mantles of hot stars because the words "chromosphere" and "corona" have very specific meanings in the context of the Sun. It is straightforward to deduce that much of what is seen to occur in the mantles of hot stars is the result of the deposition of nonradiative energy and momentum in the outer parts of the atmospheres of hot stars, but it is not easy to make successful models of the mantles of hot stars.

II. PHOTOSPHERES OF O AND WOLF-RAYET STARS

The photosphere of a star corresponds to the boundary layer required by the theory of stellar structure and evolution. Traditionally, the properties of a photosphere are determined by the composition of the atmosphere, by the value of the effective temperature of the star, and by the value of the acceleration of gravity at the surface of the star. The photosphere lies between the level in the atmosphere of the star where the gas of the atmosphere becomes opaque in all continuum wavelengths and some higher level where the gas is opaque only in the wavelengths of the strong absorption lines. The deposition of nonradiative energy and momentum appears to be unimportant for determining the physical state of the photosphere.

One of the best features to analyze for finding the properties of the photosphere is the continuous spectrum over a wide wavelength range. Because O and Wolf-Rayet stars generally are distant from the Earth, and because they usually lie in the plane of the Galaxy behind interstellar dust and gas, it is necessary to correct the observed energy distributions of O and Wolf-Rayet stars for wavelength-dependent interstellar extinction before comparing the observed energy distributions with predicted ones. A first step in obtaining observed energy distributions is, of course, the correction of the recorded signal in each wavelength band for the wavelength-dependent sensitivity of the observing instrument and the detector, and for the wavelength- and time-dependent transmission of the Earth's atmosphere.

Experience has shown (Underhill et al., 1979; Underhill, 1982) that the shape of the continuous spectrum from about 1200 to about 10000 Å of O stars and the total amount of energy radiated give useful information about the effective temperature of the star. At wavelengths longer than about 1 μm, some O stars are known to show more flux than that expected from a normal photosphere (Barlow and Cohen, 1977; Castor and Simon, 1983). The excess infrared radiation is believed to be radiated by low-density gas around the star. Analysis of the infrared excess may lead to information about the mantle of the star.

Many broad emission lines and a few absorption lines occur throughout the spectrum of a Wolf-Rayet star. Consequently, it is difficult to determine accurately the shape of the continuous spectrum of a Wolf-Rayet star. Nevertheless, some information about conditions in the photospheres of Wolf-Rayet stars has been obtained from analysis of the continuous spectra of Wolf-Rayet stars (Underhill, 1980a, 1981, 1983b; Nussbaumer et al., 1982). Most Wolf-Rayet stars show strong infrared excesses (Hackwell et al., 1974; Cohen et al., 1975; Underhill, 1980a).

Analysis of absorption lines of moderate or weak strength in the spectra of O stars may lead to information about the photosphere or about the mantle. If one deduces the same $T_{\rm eff}$ and $\log g$ from the analysis of some absorption lines as one finds from an analysis of the continuous spectrum and the wings of Hγ and Hδ, one may infer that the concept of a mantle is redundant for interpreting those lines. However, if the results of analyzing the continuous spectrum and some of the lines of O-type spectra indicate the need for models having different properties, one must retain the concept of a mantle. The properties of the photosphere will be the same as those of the model atmosphere which enables one to interpret most of the continuous spectrum successfully.

Since traditional, classical model atmospheres do not predict the occurrence of emission lines, it is clear that the analysis of the emission lines which occur in the spectra of Of and Wolf-Rayet stars will lead to information about the properties of the mantles of these types of star.

The typical depth in the atmosphere at which an absorption line is formed compared to the depth at which the continuum is formed depends on the value of $\ell_\nu/(\varkappa_\nu + \sigma)$ at the center of the line, as well as on the broadening function for the line. Here ℓ_ν is the line absorption coefficient at the center of the line, $\varkappa_\nu$ is the continuous absorption coefficient at ν, and σ is the electron scattering coefficient. If the typical value for $\ell_\nu/(\varkappa_\nu + \sigma)$ in the stellar atmosphere is small, the line will probably be formed in the same layers in which the continuum is formed; namely, in the photosphere. If it is large, the line will probably be formed in the mantle. At the low densities and high electron temperatures of O-type atmospheres, electron scattering is the dominant source of continuous opacity in the visible spectral range. Typically, $(\varkappa_\nu + \sigma)$ may have a value near 0.35 cm^{-2} per gram of star material. Consequently, most absorption lines in the spectra of O stars have values of ℓ_ν such that they are formed in layers of the atmosphere well outside the layers in which the continuous spectrum is formed.

III. MANTLES OF O AND WOLF-RAYET STARS

Modern observations of O and Wolf-Rayet stars have revealed six phenomena which cannot be interpreted by traditional, classical procedures for analyzing stellar spectra. They are:

1. The emission of X rays (see, for instance, the discovery papers by Harnden et al., 1979 and Seward et al., 1979, as well as Sanders et al., 1982)

2. The presence of emission lines in the spectra of O and Wolf-Rayet stars

3. The presence of an infrared excess

4. The presence of outflow at supersonic speeds

5. The presence of higher ions in the atmosphere of the star than can be readily accounted for using the concept of radiative equilibrium and the known effective temperatures of the O and Wolf-Rayet stars

6. The presence of inhomogeneities in the outer atmospheres of O stars

None of the above phenomena can be explained by means of normal, classical model atmospheres composed of plane parallel layers of gas in hydrostatic and radiative equilibrium. Use of the concept of statistical equilibrium (non-LTE) in place of local thermodynamic equilibrium (LTE) for calculating the distribution of the atoms and ions among their possible energy states does not solve the problems presented by the observations of the spectra of O and Wolf-Rayet stars.

To interpret what is seen, one must postulate that high electron temperatures, of the order of 10^6 to 10^7 K, occur somewhere in the outer atmosphere, that the outer atmosphere is extended, that at least some of the gas in it has been accelerated to high speeds of outflow, and that the outer atmosphere is inhomogeneous, its properties not being symmetrical in spherical shells. All of these requirements are in conflict with the requirements which define models for the photospheres of O and Wolf-Rayet stars. The traditional model atmosphere of plane parallel layers in hydrostatic and radiative equilibrium is satisfactory for explaining most of what is seen in light which escapes from the photospheres of O and Wolf-Rayet stars, but it does not serve to explain features seen in strong lines and in opaque parts of the continuum.

Because a model atmosphere, which will represent well the formation of spectral features shown by light which escapes from the outermost levels of an O or Wolf-Rayet star, must be developed using different concepts than those required for developing a model of the photosphere, we have suggested that the outer part of the atmosphere of a star be given a separate name—mantle.

The spectroscopic features which seem to be giving information about conditions in the mantles of O, Wolf-Rayet, and hot subluminous stars are discussed in Parts I and II of this book. In the next chapter, the existing theories of stellar atmospheres and of stellar spectra relevant for hot stars are reviewed, including those parts of the theory which refer to spectrum formation in extended moving atmospheres. Also, the deductions about the physical state of the atmosphere which can be made by comparing observed and predicted spectra of O and Wolf-Rayet stars are brought together. In Chapter 8, an attempt is made to establish the physics that one must use to make a successful model of the mantle of an O or a Wolf-Rayet star.

The equations which govern the determination of the temperature, density, and state of motion in the photospheres and mantles of O and Wolf-Rayet stars are derived from the conservation laws for energy, mass, and momentum, assuming the equation of state for a perfect gas. In the case of models of the photospheres of single isolated stars, it is postulated that the radiation field emergent from the envelope of the star impinges on the bottom layers of the model photosphere and that no radiation or energy in any other form is incident from the outside. Energy is transmitted through the model photosphere in such a way that radiative equilibrium is maintained. In addition, it is postulated that the gas is in hydrostatic equilibrium under the attraction of gravity due to the mass of the star. The model is determined by solving the conservation equations under these constraints. The free parameters are the composition of the photosphere, T_{eff}, and log g.

Models for mantles should be constructed after taking into account the facts that: (1) energy from a source which is not tapped in the photosphere is transmitted to the gas, heating it, and (2) at least some of the gas is flowing outward from the star at velocities sufficiently great for the gas to escape. At this time, there is little consensus of opinion about what is the hitherto unconsidered source of energy which is tapped, and about what is the mechanism by which kinetic energy is imparted to the gas and a sufficient outward impulse is given to at least some of the gas that the gas may escape as a wind.

In Part III, we shall consider whether the energy released by the action of magnetic fields is an appropriate source for the needed nonradiative energy, and whether magneto-dynamic effects in a high-temperature plasma are important for creating conditions like those we infer to be present in the mantles of O and Wolf-Rayet stars. This direction of enquiry follows naturally after one considers what appears to occur in the mantle of the Sun (i.e., in the solar chromosphere, corona, and wind).

Suitable parameters for defining models of the mantles of hot stars may be: (1) a number typifying the energy density of the source of energy which is tapped, and (2) a number characterizing the efficiency of the process which transfers energy from its storage mode into the gas in the mantles of O and Wolf-Rayet stars.

Questions requiring answers include the following:

1. What is the source for the nonradiative (nonthermal) energy and momentum which are released and cause a mantle to form?

2. By what means are the newly released energy and momentum transferred to the gas in the mantle?

3. Why is the release of nonradiative energy and momentum not significant in the photosphere?

4. How important are nonradiative phenomena in stellar atmospheres for understanding the evolution of stars?

5. Do the observations of nonradiative phenomena in stellar atmospheres give guidance on determining the stage of evolution of an isolated single star?

The stage of evolution of a star is determined by the manner in which a star generates energy in its core and by how efficiently it transfers energy through its envelope to the boundary layer or atmosphere. The boundary layer is the place where the transmitted energy can be observed. We must ask whether the detection of nonradiative phenomena in a stellar atmosphere gives an unambiguous indication of what is occurring in the interior of the star or whether the only clues to the stage of evolution are the instantaneous mass, effective temperature, and radius of the star. Since it is true that some of the spectroscopic phenomena seen for O and Wolf-Rayet stars and interpreted to mean that a mantle is present appear to be the same for young massive stars as for old low-mass hot stars in late stages of evolution, it seems that the occurrence of nonradiative phenomena in the atmospheres of hot stars is not an unambiguous indication of the stage of evolution. Information about the mass and radius of the star is needed in addition to spectroscopic information if one is to deduce the stage of evolution of the star. In the case of most single hot stars, one cannot accurately deduce masses solely from the details of the spectrum of the star. Masses can be deduced, however, when one has double-lined spectra of eclipsing binaries (see Part I).

7

MODEL ATMOSPHERES AND THE THEORY OF SPECTRA FOR O AND WOLF-RAYET STARS

I. INTRODUCTION

The spectra of the O and Wolf-Rayet stars contain emission lines as well as absorption lines. Some O stars (for instance, the well-known star 10 Lac, O8 III, or O9 V) show only absorption lines in the visible spectrum, but many O stars (usually designated as Of stars) show a few selected emission lines as well, the emission lines being weak with rather narrow cores. In particular, the stars of earliest spectral type, types O3, O4, and O5, nearly always show emission lines in addition to the absorption-line spectrum by which they are classified. The spectra of the Wolf-Rayet stars are dominated by broad emission lines, but they also contain a few intrinsic absorption lines. Because of this mixture of absorption and emission lines, the spectra of O and Wolf-Rayet stars are difficult to interpret. Throughout Part III we discuss the interpretation of the spectra of single stars only.

The classical theory for interpreting the spectra of hot stars is based on modeling processes that are appropriate for interpreting absorption-line spectra. Such theories can be used for interpreting the spectra of the absorption-line O stars. However, the models and the modeling process must be radically changed if the strengths and shapes of the emission lines of O and Wolf-Rayet stars are to be interpreted.

In this chapter, a brief summary of the chief points of classical spectrum analysis, as it is applied to single hot stars, will be presented. This will be followed by a report on the attempts which have been made to obtain an understanding of the meaning of emission lines in the spectrum of a single hot star. The studies of the theory of Of-type spectra which have been done (Mihalas and Hummer, 1973) give insight into the character of the physical conditions in the part of the atmosphere where the emission-line spectrum is formed, but they have given little information on the fundamental cause why some early-type stars show emission lines and some do not. The basic reason why some stars show Wolf-Rayet-type spectra is also not understood, although some understanding of what sort of physical situation is being seen has been obtained.

A. Traditional Model Atmospheres

The idea of representing the process of spectrum formation by means of analytical expressions and of using these expressions with tables giving the dependence of electron pressure, particle densities, and temperature on depth in the atmosphere of the star is almost as old as the techniques for observing stellar spectra. Predicted spectra found in this way are essential for obtaining an idea of what stellar spectra mean. A comprehensive review of the work which has been done in this field from its beginnings until about 1964 has been given by Pecker (1965). Here Pecker evaluates the numerical methods first used and he discusses the problems caused by the wavelength dependence of the opacity in stellar atmospheres.

The first truly successful attempts at modeling the atmospheres of O and early B stars are based on the methods pioneered by Strömgren (1940, 1944) in the context of F and G stars. First Rudkjøbing (1947), and then, independently, Pecker (1950) and Underhill (1950, 1951) developed model atmospheres for hot stars and predicted some of the properties of the continuous and line spectra from these models. With the advent of electronic computers, this type of work was greatly extended by Underhill (1962, 1968a, 1972) and by Mihalas (1965), both of whom published several models suitable for representing the atmospheres of O stars. A model atmosphere for the O star 10 Lac was constructed by Traving (1957) using similar procedures. Coarse analysis for eight O stars was carried out by Oke (1954) following the methods of Unsöld (1942, 1944). Peterson and Scholz (1971) analyzed the lines of six O-type stars using LTE model atmospheres, but they did not publish their model atmospheres.

In the first work on model atmospheres for hot stars, an iterative process was followed to obtain an acceptable model atmosphere which satisified the constraints defining the model atmosphere. These constraints were that radiative and hydrostatic equilibrium should exist. It was assumed that the distribution of the atoms and ions over their several energy states was that occurring in local thermodynamic equilibrium (LTE). The procedure was to adopt a preliminary temperature law giving the dependence of temperature on depth and then to find a model atmosphere which was in hydrostatic equilibrium (i.e., a grid of values of density, temperature, and absorption coefficients as functions of depth in the model was prepared). The radiation field was then calculated and tested to determine if the constraint of radiative equilibrium was met throughout the model atmosphere. If necessary, the temperature law was modified, and the process was repeated until satisfactory results were obtained. (See any of the papers quoted above for details). The parameters defining a traditional model atmosphere are the composition, T_{eff}, and log g. The first models for O and B stars were assumed to be composed only of hydrogen and helium.

In the case of O stars, the electron temperatures are high, being greater than 20000 K throughout the model atmosphere, and electron scattering is an important source of opacity in much of the model atmosphere. This makes the procedures used in the early work for calculating the radiation field and determining the appropriate temperature law converge slowly. More efficient methods of calculation, suited to large computers, have been developed by Auer and Mihalas (1969, 1972), Mihalas et al. (1975), and Auer and Heasley (1976).

In their complete-linearization method, Auer and Mihalas apply the powerful numerical procedures suggested by Feautrier (1967, 1968) for solving simultaneously the equations defining the structure of the model atmosphere and those defining the transfer of radiation throughout the model. Use of this method makes it possible to use the condition of statistical equilibrium for determining the distribution of the atoms and ions over their several energy states instead of assuming LTE. A set of realistic model atmospheres for hot stars, produced in this way, has been published by Mihalas (1972b). Some are suitable for representing the atmospheres of O stars. In the Mihalas model atmospheres, line-blanketing from the hydrogen lines is taken into account. Continuous absorption from a fictitious light element is introduced to simulate the effects of absorption in the continua of the C, N, and O atoms and ions. These methods have been applied by Kudritzki (1976), Kudritzki and Simon (1978), and Hunger et. al (1981) to make model atmospheres for hot subluminous stars. Their model atmospheres are composed of hydrogen and helium only.

Detailed predictions of the continuous spectra from fully line-blanketed model atmospheres have been provided by Kurucz (1979). He used numerical methods developed by himself, Avrett, and Peytremann and assumed LTE for calculating the distribution of the atoms and ions over their several energy states.

At this time quite a few model atmospheres and sets of predicted spectra exist with which to compare the observed spectra of O stars. However, little comparable work exists with which to compare the observed spectra of Wolf-Rayet stars. Presumably, the predicted continuous spectra from model atmospheres made to represent the atmospheres of O and B stars will serve as a first approximation to the *continua* of Wolf-Rayet stars.

The equations describing the variation with geometrical distance from some reference level of temperature and pressure in a traditional, classical model atmosphere are derived from the conservation laws for mass, energy, and momentum and the equation of state for a perfect gas. In the traditional model atmospheres for O stars, the geometrical configuration is plane parallel layers; the gas is assumed to be in hydrostatic-equilibrium under the inward-directed attraction of gravity (assumed constant and equal to GM_*/R_*^2 and the outward-directed force due to radiation pressure. No energy is incident on the model atmosphere except that injected at the bottom of the atmosphere by the radiation field coming from the core of the star. The amount of energy transferred across each square cm of surface is described by the parameter effective temperature. Energy is transported only by radiation.

Since the atmosphere is assumed to be in radiative equilibrium, the amount of radiant energy transferred across each square cm of surface must be constant throughout the atmosphere. At all depths in the atmosphere, the following relation is valid:

$$\sigma_R \, T_{eff}^4 = \pi \int_0^\infty F_\nu(z)\, d\nu. \qquad (7\text{-}1)$$

Here σ_R is the Stephan-Boltzmann constant and $F_\nu(z)$ is the function known as the monochromatic net flux in the radiation field at depth z. The ***true*** net flux in erg cm^{-2} s^{-1} Hz^{-1} is πF_ν. Enforcing Equation (1) constrains the temperature law to its desired form.

If LTE is assumed, the temperature usually decreases outward. Under some circumstances, a small rise may occur in the outer part of the atmosphere (Dumont and Heidmann, 1973). If non-LTE physics is used, there is a minimum temperature in the outer layers of the model atmosphere, and then the temperature rises a little as the distance from the center of the star increases. The rise of the temperature in the outermost layers is always small; the electron temperature here never exceeds the effective temperature of the star. See *Models of Static Plane Parallel Layers and Their Spectra* for some illustrative results. This effect has been discussed by Menzel, Aller, and Baker (1938) and by Cayrel (1963).

Emission lines are not usually produced by traditional model atmospheres. The spectra from traditional model atmospheres consist of absorption lines and continua with discontinuities at the ionization edges of the main constituents. The temperature rise and departures from LTE populations for the energy levels of the constituents in the outermost layers of traditional model atmospheres which have been constructed using non-LTE physics result in a strong absorption core for the profiles of the leading members of the H I, He I, and He II series (Mihalas, 1972a; Auer and Mihalas, 1972). The presence of these cores increases the equivalent widths of the H I, He I, and He II lines over what they are in the case that the condition of LTE is assumed to be valid. The wings of the Stark-broadened lines of H I, He I, and He II are almost unchanged by using the techniques of non-LTE in the place of assuming LTE.

B. Nontraditional Model Atmospheres (Models for Mantles)

To interpret emission features in a stellar spectrum, one requires a nontraditional model atmosphere (i.e., a model for a mantle). There are three ways in which the theory for predicting the spectra of single stars may be modified so that emission lines and emission continua result. They are the following:

1. The atmosphere may be postulated to have spherical geometry and the distribution of temperature and particle densities in the atmosphere to be such that the star appears bigger when viewed in the wavelengths of the emission feature than when viewed in wavelengths from the nearby continuum.

2. The electron temperature may increase outwards by a significant amount, making the source function in opaque parts of the spectrum higher than it is in the transparent parts of the spectrum which define the level of the reference continuous spectrum.

3. Emission in a few selected lines may be generated by particular radiative processes which feed atoms or ions preferentially into the upper levels of a few spectral lines. Thereby, an increased value is generated for the source function of these lines only.

Each of the foregoing changes may be brought about by some departure from the traditional set of postulates for making a model atmosphere and predicting the spectrum which it will emit.

Modification 1 is the sort of thing that was envisaged by Beals (1929) and by Menzel (1929) when they proposed the first models for accounting for the emission-line spectra of Wolf-Rayet stars. Modification 2 can give emission lines and continua with a geometry of plane parallel layers. It will result in a general emission-line spectrum, particularly if its effects are reinforced by an extended geometry as envisaged in modification 1. The type of process described in modification 3 causes the lines of multiplet 2 of N III at 4634, 4640, and 4641 Å to be in emission in Of stars (Mihalas and Hummer, 1973). Modification 3 is invoked to account for the appearance of selected lines in emission in the spectra of many types of stars (see, for instance, Bowen, 1947) and for the fact that some lines in the solar spectrum go into emission *before* the line of sight crosses the limb of the Sun when going from the disk to the chromosphere (see Canfield, 1969). Because some special relationship exists between the energy levels and transition probabilities of a minor constituent of the model atmosphere, a few lines appear in emission. Exactly which lines appear in emission depends on the local conditions. These special effects cannot occur unless the local situation departs from the state of thermodynamic equilibrium.

To make a model for a mantle, one should work from the conservation laws for energy, momentum, and mass, using the equation of state for a perfect gas and postulating a suitable geometry. The energy incident on the model mantle comes from the radiation field from the photosphere (calculated by means of a traditional model atmosphere) and from any nonradiative source of energy which deposits energy in the mantle. Possible sources for nonradiative energy are wave motions generated in the envelope of the star, the rotation of the star, and the ambient magnetic fields. To determine the temperature law in the model mantle, the gains from these three sources of energy are compared with the losses due to the escape of photons and of gas, in the form of a wind, from each element of volume in the mantle.

To find the density structure in the model mantle, the equations of conservation of mass and of momentum are considered. Account must be taken of the gains in momentum from the action of accelerating forces such as radiation pressure and of the outward-directed forces resulting from magnetohydrodynamic effects, as well as the losses due to the action of gravity and of inward-directed magnetohydrodynamic forces. Electrostatic forces may be significant in some cases. Boundary conditions may be set by requiring that the rate of mass loss have a prescribed value and that the pressure and density structure be continuous with that in the model photosphere which underlies the model mantle. At a great distance the mantle must meld with the interstellar medium.

The existing models for mantles, particularly those used in the theories of line formation in moving extended atmospheres, are described by means of ad hoc velocity laws, spherical geometry, and the condition that the rate of mass loss be equal to a prescribed value. The adopted temperature law is ad hoc in all cases. The existing theory for line formation in winds is not internally consistent in the sense of solving simultaneously all three conservation equations to determine local values for temperature, density, and flow velocity.

C. Plan of this Chapter

In *Models of Static Plane Parallel Layers and Their Spectra*, the results obtained using traditional, plane parallel layer model atmospheres for hot stars are reviewed. Some predicted line profiles are compared with the observed line profiles in a normal O star, and the deductions which have been made about the physical state and composition of O-type atmospheres are summarized.

In *The Continuous Spectrum from Spherical Model Atmospheres*, the methods for making spherical model atmospheres are summarized, and the predicted continuous spectra from spherical model atmospheres are compared with those from plane parallel layer, static model atmospheres and with the observed continuous spectra of O stars. The theory of stellar winds seen by means of the continuous spectrum is summarized. How the rate of mass loss from a star may be estimated from observations of part of the continuous spectrum of the star is reviewed.

In *The Line Spectrum from Moving Three-Dimensional Model Atmospheres*, the existing theory of the formation of lines in moving extended atmospheres is reviewed, and typical predicted line profiles are compared with observed line profiles in the spectra of O and Wolf-Rayet stars. The theory of line-driven winds is summarized, as well as the attempts to determine the composition and physical state of the atmospheres of Wolf-Rayet stars.

The methods for estimating the rate of mass loss from hot stars by means of analyses of the spectrum are reported in *The Rate of Mass Loss from Hot Stars*, and a summary of the information available for O and Wolf-Rayet stars is given.

Conditions in flowing model mantles may be such that plasma instabilities occur. Knowledge on this subject, as it is applied to the atmospheres of hot stars, is reviewed in *Atmospheric Stability: X Rays*. High temperatures may be generated in some parcels of gas leading to the emission of X rays.

II. MODELS OF STATIC PLANE PARALLEL LAYERS AND THEIR SPECTRA

The simplest way to model stellar atmospheres is to consider what may occur as radiation passes through plane parallel layers of gas.

A. Methods for Making Model Atmospheres

The methods for making traditional model atmospheres defined by the constraints of radiative and hydrostatic equilibrium have been outlined above. One assumes that the atmosphere is confined in plane parallel layers of gas, and that the density and temperature at each place in the atmosphere are functions of one geometric coordinate only, z, the distance along the line of sight measured from the center of the star outward. A composition giving the relative numbers of each element present (usually only hydrogen, helium, and sometimes a few light elements are considered), a value of log g, and a temperature by which to identify the model atmosphere are adopted. One then proceeds to find density and temperature as functions of the geometric depth in the model, these quantities being so related that the perfect gas law is satisfied and the temperature and pressure gradients are such that radiative and hydrostatic equilibrium exist.

At the outer boundary of the model, the pressure is made to approach zero; the temperature approaches a boundary value, which is determined by the adopted temperature which identifies the model, and the geometry. In the case that local thermodynamic equilibrium (LTE) is assumed for finding the distribution of the ions and atoms over their several energy states (i.e., for finding the partition function for each species), the boundary temperature approaches a value in the range 0.7 ± 0.08 times T_{eff} for models of O and early B stars (see Figure 15 of Pecker, 1965). In the case that the condition of statistical equilibrium (non-LTE) is used to determine the partition functions, the temperature rises somewhat from these values, but remains less than the effective temperature of the model (Mihalas, 1972b). These statements are illustrated below.

The methods first followed in making model atmospheres consist of assuming a preliminary temperature law, usually that for a gray body, and solving the equation of hydrostatic equilibrium to find a preliminary model atmosphere. Then the radiation field of this model is calculated and tested to see if radiative equilibrium exists. If radiative equilibrium does not exist, the temperature law is changed, and the process is repeated until the departures from radiative equilibrium are sufficiently small that they may be neglected.

One requires a formula telling *how* to change the temperature law and an understanding of *how closely* radiative equilibrium, that is, constant net flux at all levels in the model (see Equation (7-1)), must be enforced in order that a model atmosphere may be an acceptable representation of a real stellar atmosphere. The accuracy with which the constraint of hydrostatic equilibrium is enforced is simply a matter of calculating techniques and of how finely the depth parameter is divided. These topics have been discussed exhaustively in the literature. When the absorption coefficients, as functions of wavelength or frequency, vary over a wide range of values and electron scattering occurs, calculating the radiation field accurately presents very real problems. A review of the findings concerning determining the temperature law is given in Pecker (1965), in Section 6-3 of Mihalas (1970), and in Section 7-2 of Mihalas (1978).

When making some of the first O-type model atmospheres, Underhill (1951, 1968a), chose to find an acceptable temperature law by trial and error. Mihalas (1965), in his set of LTE model atmospheres, applied temperature-correction procedures based on the suggestions of Avrett and Krook (1963a, 1963b). When electron scattering provides a significant part of the opacity in the atmosphere, temperature correction methods based on the assumption that the source function be approximated by the Planck function do not work well.

In all the early work, the starting temperature law is described in terms of an optical depth measured by an absorption coefficient that is gray, and thus has the same value at all wavelengths. There are several ways of defining the gray absorption coefficient (see, for instance, the reviews by Pecker, 1965, and by Mihalas, 1970, 1978). Each investigator has selected the definition which will give desired advantages in some part of the process of computing and testing the radiation field in the model atmosphere. The linear depth is replaced as the significant depth parameter for the model atmosphere by the optical depth, τ, corresponding to the adopted gray absorption coefficient. We have

$$\tau = - \int_z^{\infty} \kappa \rho \, dz. \qquad (7\text{-}2)$$

Here κ represents the gray absorption coefficient and ϱ is the density in the atmosphere at level z.

The second way of making models to represent the atmospheres of O and early B stars is to solve the equations of hydrostatic equilibrium and radiative equilibrium simultaneously by using the scheme proposed by Feautrier (1967, 1968). Then the concept of statistical equilibrium (non-LTE) may be used to find the occupation numbers for the various energy states of the constituents of the model atmosphere. Using the concept of statistical equilibrium introduces a major difficulty because the occupation numbers in the outer parts of the model are sensitively dependent on the detailed structure of the radiation field. A powerful complete-linearization method to solve such problems has been developed by Auer and Mihalas (1969, 1972) and improved by Auer and Heasley (1976). The concepts which are used are described in Section 7-5 of Mihalas (1978). Usually the radiation field and particle densities of an LTE model atmosphere are used as the starting model. When non-LTE model atmospheres are made, the geometric depth parameter, z, is replaced by a parameter which gives the amount of mass in the column above the level in question. Thus, the variable z is replaced by

$$m = -\int_z^{\infty} \rho \, dz. \qquad (7\text{-}3)$$

The relationship between m and τ depends on how the gray absorption coefficient $\varkappa$ varies with linear depth, z, in the model atmosphere.

In the first method used to make model atmospheres for hot stars, the equations of hydrostatic and radiative equilibrium are solved in series. The techniques used to compute the radiation field are optimized to make the procedure converge rapidly. In the second method, the equations of hydrostatic and radiative equilibrium are solved in parallel, appropriate numerical methods being used to obtain rapid convergence of the procedure. This method requires a computer with a large rapid-access memory because large matrices of data must be manipulated.

Physics. A major part of making model atmospheres and predicting the spectrum that they will emit is concerned with computing the radiation field in the model atmosphere. This means that one must have expressions for the values of the line and continuum absorption coefficients of all the atomic and ionic species present in the atmosphere. These coefficients are written in terms of the local temperature and electron density, as well as in terms of the wavelength or frequency of the radiation. One also needs cross sections for transitions resulting from collisions with electrons.

Obtaining such information from the theory of the interactions between a gas and radiation and from laboratory experiments has been an important part of the work on model atmospheres and the prediction of spectra. The required expressions can be found in the major papers in this field and in textbooks dealing with the physics of stellar atmospheres. One useful source for the first expressions used is Unsöld (1955); recent work, relevant for making models of the atmospheres of hot stars, is summarized in Chapter 4 of Mihalas (1970, 1978). Other textbooks on interpreting stellar spectra also review the basic knowledge of the physics of gases and radiation that is needed.

When it is desired to predict line profiles, particular attention must be paid to evaluating the shape of the line-absorption coefficient, as well as to determining the intrinsic probability for the transition of interest to take place. In the work on stellar spectra, transition probabilities are generally used in the form of gf values.

In the atmospheres of hot stars, the broadening processes due to Stark effect, to radiation and collision damping, and to thermal motion must be evaluated. The broadening functions are expressed in terms of the local temperature and the local density of particles at each place in the model atmosphere. In addition, one must consider the appearance of the predicted spectrum as it will be seen by a distant observer. Since each observation records light from all parts of the disk of the star, one must allow for the effects of the rotation of the star and of macroturbulence. Macroturbulence is the term used to represent the broadening of a line profile, seen by a distant observer, which results from the displacements caused by randomly distributed in-and-out movements of parts of the stellar atmosphere.

Motion along the line of sight displaces the spectral lines, motion toward the observer resulting in a displacement to shorter wavelengths. Broadening due to rotation of the star and to macroturbulence is geometrical in origin. It is due to the fact that normally in stellar spectra the light from the full disk of the star is recorded at one time. A useful review of rotation and macroturbulence as they affect stellar spectra has been given by Huang and Struve (1960).

Another broadening agent, microturbulence, is often discussed in connection with the interpretation of the profiles and equivalent widths of stellar absorption lines. Microturbulence is a *fictitious* broadening agent which was first introduced (Struve and Elvey, 1934) to account for the fact that the strong lines of some multiplets of Si III are relatively stronger than the weak lines in the case of spectra of supergiants than for the case of main-sequence stars. Microturbulence is postulated to act on the curve of growth in the same way that an increased thermal Doppler broadening would act. (The curve of growth tells how the equivalent width of a line increases as the number of atoms or ions absorbing the line increases along the line of sight. Curve-of-growth analyses of multiplets of lines use a model consisting of a single layer of gas at one temperature and one pressure to interpret the observed equivalent width of each line and to find the temperature, pressure, and composition of the layer of gas.)

However, it is postulated that microturbulence does not alter the relative degree of ionization and excitation of the species giving the spectrum. The partition function for the species is calculated as though LTE exists. Just as for thermal Doppler broadening, the presence of "microturbulence" results in broadened spectral lines. Modern work, in particular that of Mihalas and his colleagues using non-LTE physics, has shown that most of the effects attributed to "microturbulence" in the atmospheres of hot stars are really due to the increased strength for some lines which results from using a non-LTE theory of spectrum-line formation. Microturbulence is a fictitious factor which is introduced into LTE analyses of stellar spectra to make intrinsically weak lines and intrinsically strong lines yield about the same abundances. (See Chapter 6 of Underhill and Doazan (1982) for a list of spectra which have been studied by non-LTE physics in hot model atmospheres.)

Information about representations for continuous and line absorption coefficients can be found in most textbooks on the physics of gas and radiation and on spectroscopy, as well as in most textbooks about interpreting stellar spectra. Information relevant to interpreting the spectra of O and Wolf-Rayet stars can be found in most of the papers referenced below in our discussion of the composition and physical state of the atmospheres of O and Wolf-Rayet stars.

1. Identifying a Model Atmosphere with a Real O Star. A model atmosphere is identified with a real star by demonstrating that the spectrum from the model atmosphere reproduces major features in the spectrum of the star in an acceptable manner. When deciding whether the agreement between prediction and observation is acceptable or not, one must consider two questions. The first is whether or not the expressions which have been used to evaluate the transfer of radiation through the model atmosphere are sufficiently realistic to generate confidence that the model does, indeed, represent well the physical state in the atmosphere of a real star. The second is whether or not the constraints of radiative and hydrostatic equilibrium which have been used to fix the range of temperature and density which are encountered in the observable parts of the atmosphere are met with sufficient accuracy that the apparent agreements and discrepancies between the predicted spectrum and an observed spectrum are meaningful. In most of the work on O stars to be reported below, the comparison between observed and predicted spectra has been carried out using a few spectral features in the part of the spectrum accessible from the surface of the Earth (i.e., the visible spectrum). It is relevant to inquire whether the criteria that have been used select a model for the photosphere or whether they select a model for a part of the mantle of the star. The answer to this question is not immediately obvious, as the following examples will show.

A traditional model atmosphere is a tabulation of temperatures and particle densities as a function of depth in plane parallel layers of gas of a fixed composition. The spectrum which the model emits is described by the wavelength dependence of the net monochromatic flux of radiation evaluated in the outermost layer of the model.

Formally, the net monochromatic flux in any layer of the model lying at an optical depth, τ, can be found by evaluating the expression

$$F_\nu(\tau) = 2\left[\int_\tau^\infty S_\nu(t_\nu)\, E_2\,(t_\nu + \tau)\, dt_\nu - \int_0^\tau S_\nu(t_\nu)\, E_2\,(\tau - t_\nu)\, dt_\nu\right]. \tag{7-4}$$

Here τ is the gray-opacity optical depth of the layer (see Equation (7-2)), t_ν is the monochromatic optical depth at frequency ν, and E_2 is the second exponential integral. The function $S_\nu(t_\nu)$ is the monochromatic source function. It is formally defined as the emissivity at frequency ν divided by the total opacity at frequency ν. The derivation of Equation (7-4) can be found in most textbooks on radiative transfer in stellar atmospheres.

Various forms are written for S_ν in terms of the temperature, the particle densities, and the mean intensity of the radiation field. The precise form depends on the details of *how* the transfer of radiation is postulated to occur. The transfer of radiation through a stellar atmosphere is the result of interactions between the constituents of the model atmosphere and photons. Typical forms for the source function for lines and for continua are described in all research papers and textbooks on the theory of stellar spectra.

In the case of LTE, one can use thermal equilibrium relationships between the occupation numbers for the various energy states of the atoms and ions to simplify the expression for the source function. In the case of non-LTE, explicit relationships between the energy-level populations and the radiation field are taken into account. The source function is evaluated in terms of model atoms which are defined by means of sets of energy levels and the probabilities for radiative and collisional transitions in and out of these levels. Continua of energy levels, as well as energy levels having discrete energies, are considered.

Let us use a simple example to show how different parts of the spectrum of a star may be sensitive to conditions in different parts of the atmosphere. We note that the spectrum emitted by a model atmosphere is found by evaluating Equation (7-4) at a number of frequencies in lines and in the continuum for the case that $\tau = 0$. When a quadrature formula is applied to evaluate the integrals in Equation (7-4) (see, for instance, Chandrasekhar, 1950), we can write

$$F_\nu(0) = \sum_{i=1}^{n} a_i S_\nu(t_i). \tag{7-5}$$

Here the t_i are the values of t_ν at which the source function is to be evaluated and the a_i are appropriate weights. If three points are used, suitable t_i are 0.287, 1.814, and 5.385; the corresponding a_i are 0.7669, 0.2265, and 0.00659, respectively. These numbers show that the value of the emergent flux, $F_\nu(0)$, is chiefly determined by the value of the monochromatic source function at t_ν = 0.287 and 1.814.

What the value of S_ν will be for any frequency ν at the representative depths will depend on how t_ν is related to τ, and thus to temperature and density. We have

$$t_\nu = \int_0^\tau \frac{(\kappa_\nu + \ell_\nu + \sigma)}{\kappa} \, d\tau, \qquad (7\text{-}6)$$

where κ_ν is the monochromatic continuous absorption coefficient due to all sources of opacity, ℓ_ν is the monochromatic line-absorption coefficient, and σ is the coefficient for electron scattering. Here κ is the gray opacity coefficient. Depending on whether one is considering radiation at frequencies in the continuum or in lines, and on how κ is defined, t_ν may increase slowly or rapidly with increasing depth in the model atmosphere.

In opaque parts of the spectrum, such as wavelengths in the centers of intrinsically strong lines from abundant species and in the Lyman continuum of hydrogen, the depth at which $t_\nu = 1.814$ may be reached at very high layers in the model. On the other hand, at frequencies in the continuous spectrum near 4500 Å or in the near infrared, $t_\nu = 1.814$ may not be reached until deep in the model atmosphere. The result of this is that the observed fluxes in wavelengths in opaque parts of the spectrum may carry information about the physical state of quite different parts of the model than do the fluxes from transparent parts of the spectrum.

If the criteria used to establish identity between a model and an O star are all based on opaque parts of the spectrum, they may lead to the selection of a different representative model atmosphere than would be chosen if a different set of criteria based on transparent parts of the spectrum were used. This is particularly likely to be true if the temperatures and densities in the outer parts of a real O-type atmosphere are influenced by the deposition of nonradiative energy and outward-directed momentum. The physical state of the continuum-forming part of the atmosphere probably will be determined chiefly by the amount of radiation transmitted from the center of the star and by log g.

The following example shows how differently t_ν may vary with increasing depth at different wavelengths in the spectrum of an O star. The example has been worked out using the temperatures and particle densities given by Mihalas (1972b) for two non-LTE (traditional) model atmospheres. One is designated as (30000, 4.0, NLTE/L) and the other as (45000, 4.5, NLTE/L). Here the first number gives T_{eff}, the second gives log g, and NLTE/L means that the occupation numbers for the energy levels of the various constituents (hydrogen, helium, and a fictitious light element representing ions of C, N, and O) have been found by solving the equations of statistical equilibrium for the case that the lines of He II between levels with even principle quantum numbers do not overlap with the lines of H I which appear at about the same wavelengths.

The variation of t_ν with depth in the models is shown in Figures 7-1 and 7-2 for the central wavelengths of He I 4471 and He II 4541, for electron scattering, for continuous opacity in the Paschen continuum of H I at 4500 Å, and for free-free opacity at 3.6 μm. The monochromatic opacity coefficients in each layer of the models were calculated using the particle densities and temperatures given by Mihalas (1972b). To obtain an estimate of the populations of the 2^3P° levels of He I and of the $n = 4$ levels of He II, which Mihalas does not give, it was assumed that the populations of these levels may be found from the populations of the next higher ions by applying the Saha and Boltzmann laws. We have

$$t_\nu = \int_{m_0}^{m} \frac{\alpha_\nu \, N(\text{level}) \, dm}{m_A \, N(\text{H}^+)} \; . \qquad (7\text{-}7)$$

Here α_ν is the absorption coefficient per atom for the appropriate transition (i.e., for the lines He I 4471, He II 4541, the Paschen continuum at 4500 Å, and the free-free continuum at 3.6 μm). To find the opacity due to electron scattering, α_ν is replaced by the Thomson scattering coefficient per electron. The quantity N

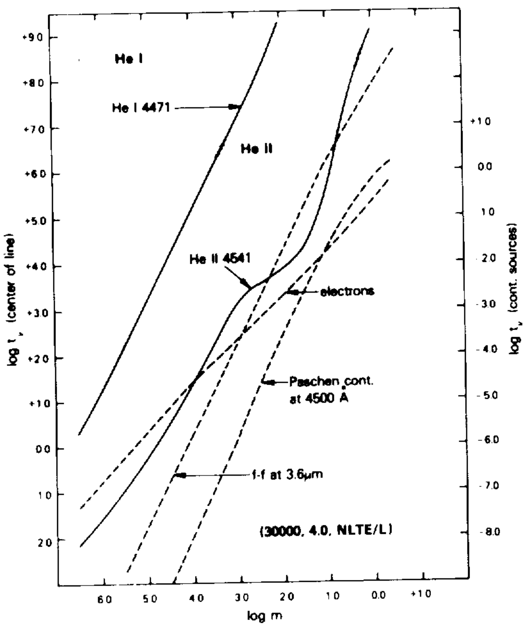


Figure 7-1. The variation of monochromatic optical depth, τ_ν, with depth in a model prepared by Auer and Mihalas (1972). The depth is measured by the quantity m, which gives the mass in grams lying above each cm^2 of surface in each layer of the model atmosphere. In this model, T_{eff} = 30000 K, log g = 4.0, and non-LTE physics is used. In the case of the lines, t_ν is calculated at the line center.

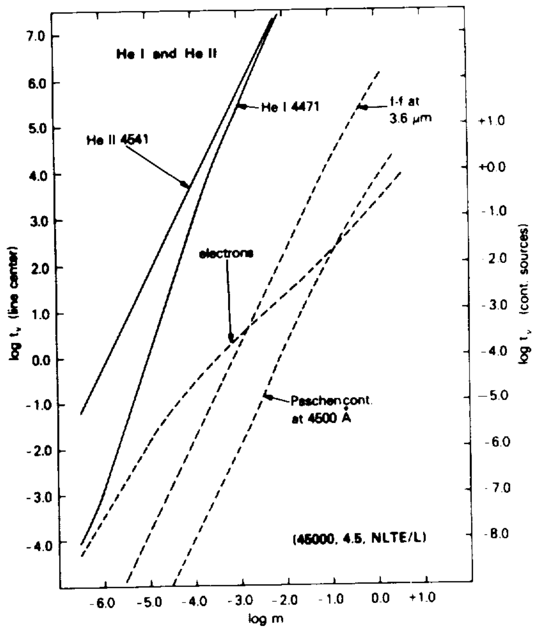


Figure 7-2. The variation of monochromatic optical depth, t_ν, with depth in a model prepared by Auer and Mihalas (1972). The coordinates are as in Figure 7-1. In this model, T_{eff} = 45000 K, log g = 4.5, and non-LTE physics is used.

(level) is the number of atoms/ions per cm^3 in the level from which absorption takes place. It is N_e, the electron density, in the case of electron scattering. The correction for stimulated emission was applied at 3.6 μm, where it is significant, but not at the other wavelengths chosen for this illustration. Neglect of this correction factor approximately takes account of the fact that there is also a little continuous opacity from He I and He II, as well as from other overlapping H I continua at the wavelengths of the lines and at 4500 Å. The line opacities are calculated for the line center. To find the value at other wavelengths, one must multiply by a normalized broadening function.

The integration is carried out inward from the first layer of the model, m_0, to the layer m where the mass of a column of gas having a cross section of 1 cm^2 is m grams. The quantity dm is the increment in mass between each layer of the model. The length in cm of each layer of the model is $dm/m_A N(H^+)$, where m_A is the average mass per particle in the atmosphere. It can be found from the information given by Mihalas about the composition of his models. The number of hydrogen atoms in each layer in all stages of excitation or ionization has been put equal to $N(H^+)$, which is a number given by Mihalas. Throughout O-type model atmospheres, the ionization of hydrogen is essentially complete.

In spite of Stark effect and thermal Doppler motion, the values of the line-absorption coefficients decrease rapidly as one considers frequencies further and further into the wings of the lines. Probably the core of the line, which is the part which is important for telling how "strong" a line appears to be on classification spectra, is formed by considering frequencies such that the line-absorption coefficient takes on values between 1 and 10^{-6} times its central value. Here we consider the part of the model atmosphere important for determining the shape and strength of the He I and the He II lines to be the region in which $0.3 < t_\nu$ (line center) $< 10^6$. The part of the model important for determining the continuum flux at 4500 Å lies in the region where t_ν (electrons) plus t_ν (H, $n = 3$) lies in the range 1.0 ± 0.8. These spectrum-forming regions are shaded gray in Figures 7-1 and 7-2. The part of the atmosphere most important for forming the continuum at 3.6 μm lies a little outside the region important for the continuum at 4500 Å, but well inside the regions important for the He I and He II lines.

In the (30000, 4.0, NLTE/L) model, the line He I 4471 is formed chiefly outside the region where the line He II 4541 is formed. Both lines are formed entirely outside the region important for determining the strength of the continuum at 4500 Å. In the (45000, 4.5, NLTE/L) model, both the He I and the He II lines are formed in the same part of the model. (One may estimate the region of formation of other He I and He II lines from the same lower levels by scaling the plotted monochromatic optical depths by the ratio of f values between the other lines and the lines shown here.) The continuum at 4500 Å is formed in a much deeper part of the model. The separation in linear distance between the region where He I 4471 is formed and the region where the continuum at 4500 Å is formed is about 2.3×10^9cm for the 30000 K model and about 1.2×10^9cm for the 45000 K model. These numerical results imply that a model which reproduces the equivalent widths of He I 4471 and He II 4541 tells us essentially nothing about the physical state of the layers in which the visible continuous spectrum is formed unless we are certain that the adopted temperature law is correct. Criteria based on the shape of the continuous spectrum must be used to select a model which will accurately represent the physical state in the deep layers of the stellar atmosphere in which the visible continuum is formed.

All of the model atmospheres now available for representing the atmospheres of O stars have been constructed using the constraint of radiative equilibrium to determine the temperature law and the constraint of hydrostatic equilibrium to determine the density structure of the model. The only input of energy is an amount of radiation corresponding to the parameter, T_{eff}. However, in the case of O stars, observations of the ultraviolet spectrum have indicated that additional energy from a nonradiative source is deposited in the outer parts of the atmosphere, heating them, and that outflow occurs. At present we do not know exactly where superheating and outflow begin, but the theoretical results shown in Figures 7-1 and 7-2 should caution us that it is possible that the model atmosphere selected by means of matching the relative strengths of He I 4471 and He II 4541 may correspond to a different, higher effective temperature than will be indicated by a model atmosphere selected by means of matching the shape of the continuous spectrum in spectral regions where the opacity is like that at 4500 Å. Which model atmosphere and T_{eff} is the *correct* one for characterizing the star?

The parameter T_{eff}, by definition (see Equation (7-1)), represents the stream of radiative energy from the star. This energy is generated by the nuclear reactions which take place in the center of the star. The model atmosphere representing the continuum-forming regions of the stellar atmosphere is the correct model to select if one wishes to determine the effective temperature of the star by means of a model atmosphere. This is because, in the deepest layers of the stellar atmosphere, nonradiative heating appears to be negligible. A model selected by means of matching the profiles and equivalent widths of strong lines in the visible part of the spectrum of an O star represents the electron temperatures and densities encountered in an outer region of the star—in a part of the mantle—where superheating may have already begun. The value of T_{eff} for a model selected by analyzing lines represents the heating caused by the radiation stream from the core of the star, as well as heating caused by nonradiative sources of energy. The difference in T_{eff} between the value for a model atmosphere which represents well the continuum-forming layers of the stellar atmosphere—the photosphere—and the value for the model which represents the layers in which the lines are formed is a measure of the amount of nonradiative energy which has gone into superheating the line-forming layers of the model atmosphere.

In Underhill (1980a), it is shown that the effective temperatures estimated for O stars from analysis of line strengths are systematically higher than the effective temperatures found from the integrated flux and angular diameters or from matching the shape of the continuous spectrum. A similar systematic effect for B stars has been reported in Chapters 3 and 4 of *B Stars With and Without Emission Lines* (Underhill and Doazan, 1982).

In the case of Wolf-Rayet stars, the line spectrum is basically an emission-line spectrum, and it clearly cannot be interpreted by a traditional model atmosphere. The effective temperature of a Wolf-Rayet star must be estimated from the integrated flux and an angular diameter for the star. Such an effective temperature may be confirmed by prediction of the shape of the continuous spectrum using a traditional model atmosphere. However, one must be alert to the possibility that the atmosphere of a Wolf-Rayet star may be so unusual that a traditional model atmosphere cannot be used to represent even the continuum-forming layers of the atmosphere accurately.

The appropriate value of log g for a model atmosphere to represent the atmosphere of an early-type star is found by matching the wings of predicted profiles of Hγ and Hδ to the observed wings for these lines. This procedure determines the electron density in moderately deep layers of the model atmosphere. By definition, $g = GM_*/R_*^2$, thus, if one knows the radius of the star from other information and can find log g from model-atmosphere analysis, one may deduce the mass of the star. This procedure is reliable only if one is sure that hydrostatic equilibrium obtains in the deep layers of the atmosphere, and if the outward-directed force of radiation pressure has been taken into account reasonably accurately in the procedures used to find the model atmosphere. Fitting the wings of the Hγ and Hδ lines determines the effective value of g in the deep layers of the atmosphere of the star. If the effects of radiation pressure or any other outward-directed force have been underestimated when making the model, the value of M_* deduced from log g for the model will be a lower limit. If the hydrogen Balmer lines are broadened by exceptionally large "microturbulence" of nonthermal origin, log g may be overestimated (Underhill, 1984b).

B. Models for O-Type Atmospheres

There are two ways of fixing the range in which temperatures in a model atmosphere will lie. One is to assign a value for T_{eff}; the other is to assign the temperature in one of the layers of the model believed to be representative for conditions where the line spectrum of a real O star is formed. The first method has been employed by most people who have made sets of model atmospheres. The second method was employed by Underhill (1950, 1951, 1962, 1968a, 1972) because, when she started making model atmospheres for hot stars, little was known about the effective temperatures of O and B stars. On the other hand, curve-of-growth-type analyses (see, for instance, Unsöld, 1942, 1944) had revealed at that time the range of temperature which is encountered in the *line-forming* layers of late O and early B stars. When the second method is used to specify the model atmosphere, the effective temperature is a result of the computations. It does not usually turn out to be exactly one of the values quoted frequently in the literature for hot stars such as 25000, 30000, and 40000 K.

A preliminary choice for log g can be made by considering the information on electron density in the atmosphere which may be deduced from the breakoff point of the Balmer series of hydrogen (Inglis and Teller, 1939). One may follow the arguments of Unsöld (1942, 1944) to estimate log g. A precise model atmosphere to represent the line-forming regions of an O star is usually found by matching the observed wings of Hγ and Hδ to find log g and by matching the equivalent widths of a few He I and He II lines to establish T_{eff}. The process of fixing log g and T_{eff} must be done iteratively, since the strength of the wings Hγ and Hδ depends to a certain extent on the adopted T_{eff}, and the computed equivalent widths of lines of He I and He II have some dependence on log g since these lines are broaded by Stark effect. Consequently, their strengths are dependent on the density of charged particles in the atmosphere, as well as on the temperatures there.

The way in which the appropriate values of T_{eff} and log g are found by model-atmosphere analysis is to find by trial and error a few pairs of values of T_{eff} and log g which will satisfy each of the criteria (i.e., fitting the profiles of the wings of Hγ or Hδ, the equivalent widths of one or two He I lines and of one or two He II lines) and plot loci for these values in a diagram of log T_{eff} plotted against log g. A different locus, often approximated by a straight line, is found for each criterion. The (log T_{eff}, log g) point defined by the intersection of several loci defines the best pair of T_{eff} and log g for a model atmosphere which will represent at one time the regions of formation of all the selected criteria.

In actual practice, the selected criteria do not always define closely a single pair of values for T_{eff} and log g. This is not surprising in view of the different parts of the model atmosphere which are sampled by each criterion (see Figures 7-1 and 7-2) and the fact that we know that heating due to the deposition of nonradiative energy and outflow begins to become apparent somewhere in the outer layers of the atmospheres of normal O stars.

A typical diagram for determining for an O star a representative pair of values for T_{eff} and log g is shown in Figure 7-3. This diagram has been taken from the model-atmosphere study of the spectra of eleven subluminous O stars by Hunger et al. (1981). Hunger and his col-

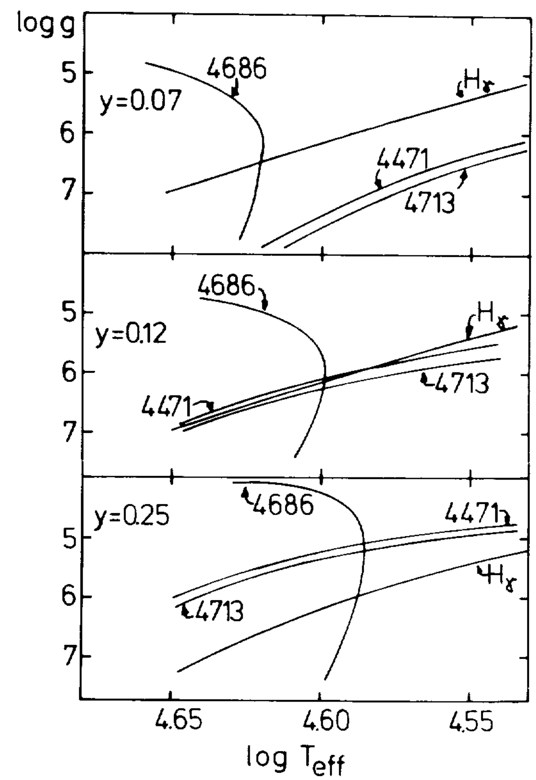


Figure 7-3. Fit of the equivalent widths of Hγ, He II 4686, He I 4471, and He I 4713 in the log g - log T_{eff} plane for a given helium abundance, y, by number (from Hunger et al., 1981).

leagues find that in some cases the strengths of the He I lines cannot be represented with reasonably good accuracy by model atmospheres of normal composition when the pair of values for T_{eff} and log g that is selected permits a reasonable representation for the strength of Hγ and the strength of He II 4686. They solve the problem by varying the composition of the model atmosphere. Changing the parameter, $y = N(\mathrm{He})/N(\mathrm{H})$, from its value for Population I stars permits them to represent all the criteria in an acceptable manner. In this way, Hunger et al, have deduced that some hot subluminous stars are underabundant in helium, y being significantly smaller than 0.1, while in other hot subluminous stars helium is overabundant, the appropriate value of y being much larger than 0.1. The value 0.1 is what has been found to be appropriate for B stars (see *B Stars With and Without Emission Lines*, Chapter 6, Underhill and Doazan, 1982). The values of T_{eff} deduced by Hunger et al. for hot subluminous stars are in reasonable accord with the effective temperatures estimated from the shape of the continuous spectrum.

Information about the first model atmospheres for O stars has been reported above. These model atmospheres have T_{eff} in the range 25000 to 50400 K, and log g = 4.0 and 4.5. They are based on the hypothesis of LTE, they are composed of hydrogen and helium only, and they do not take account of the opacity due to lines. Although these first models provided insight into where the continuous spectrum was formed in O stars, they were unsatisfactory for representing the formation of lines. (See Underhill, 1968a, for a summary of the results of LTE analysis.)

The line-blanketed model atmospheres of Kurucz (1979), although they are constructed using the hypothesis of LTE, provide realistic representations for the continuous spectra of O stars. This has been demonstrated in Underhill et al. (1979) and is illustrated in Figure 1 in Underhill (1984b). Kurucz gives model atmospheres with T_{eff} in the range 8500 to 50000 K, log g = 2.0 to 5.0. Those with effective temperatures suitable for O stars have solar composition.

To represent accurately the formation of absorption lines in the atmospheres of O stars, it is necessary to use the hypothesis of statistical equilibrium to determine the distribution of the constituents over their possible energy states. The compilation of model atmospheres by Mihalas (1972b) presents the first set of non-LTE model atmospheres for O stars. The range in effective temperature is from 15000 to 55000 K, and the range in log g is from 2.0 to 4.5. As in the case of the Kurucz model atmospheres, all values of log g are not used at every T_{eff} because hydrostatic equilibrium must be maintained against the disruptive action of radiation pressure. The chief lines of H I, He I, and He II are taken into account in the non-LTE calculation of occupation numbers. The composition approximates solar values, with y = N(He)/N(H) = 0.10 and N("light element")/N(H) = 0.0015.

The method of Auer and Mihalas (1969, 1972), see also Mihalas et al. (1975), has been used by the group at Kiel University (see, for instance, Kudritzki, 1976; Kudritzki and Simon, 1978; and Hunger et al., 1981) to make model atmospheres for hot subluminous stars. This group has explored the effects of changing the relative abundance of helium. When the abundance of helium is reduced, the lines of He I and He II are weakened, but the lines of hydrogen remain unchanged. When the abundance of helium is increased, the model atmosphere becomes more transparent in continuum wavelengths in the observable range. The lines of He I and He II increase in strength. In fact, all lines become stronger when the abundance of hydrogen is greatly reduced because of the increased transparency in continuum wavelengths. No significant change in the profiles of the Balmer lines of hydrogen is noted until $y > 0.90$. A grid of pure hydrogen model atmospheres for hot subluminous stars, most calculated in LTE, has been presented by Wesemael et al. (1980).

Typically, the electron densities and temperatures in model atmospheres which represent normal O-type photospheres depend on depth in the model as shown in Figures 7-4 and 7-5. The data to make these diagrams have been taken from the work of Mihalas (1972b). The results for the NLTE/L models are shown by solid lines, and those for LTE models by broken lines. The electron density decreases steadily outward, it being slightly larger in the case of LTE model atmospheres than for NLTE/L models at the same depth point. At depths below the point where log m = -1.0 for the model with T_{eff} = 30000 K and below the point where log m = 0.0 for the model with T_{eff} = 45000 K, the temperature in the NLTE/L model is the same as in the LTE model. Outside these depths, the temperature in each NLTE/L model rises over what it is in the LTE model. In no case does the temperature rise higher than the effective temperature.

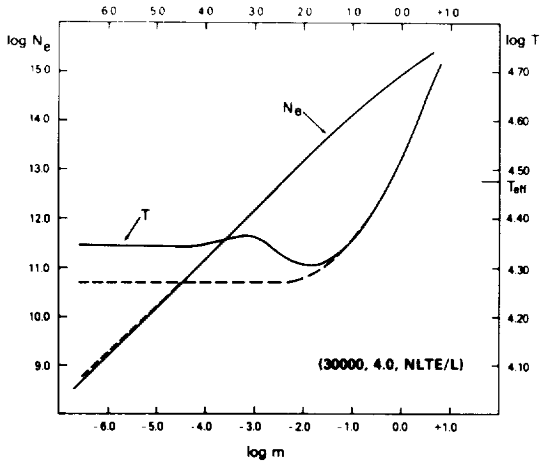


Figure 7-4. The temperature and electron-density structure in model atmospheres having T_{eff} = 30000 K and log g = 4.0. The results for LTE are shown by broken lines, those for non-LTE by full lines (from Auer and Mihalas, 1972).

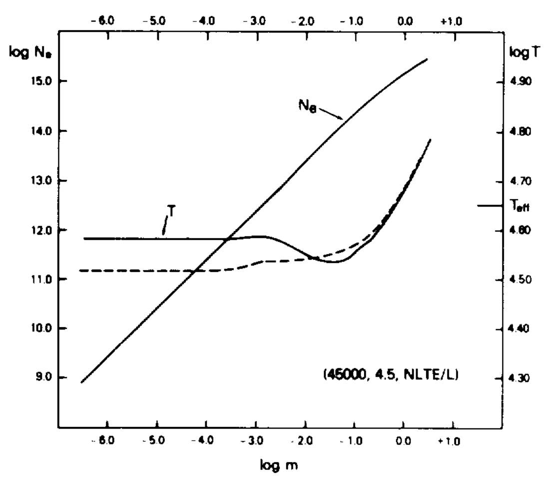


Figure 7-5. As in Figure 7-4, except that T_{eff} = 45000 K and log g = 4.5.

No matter whether the model is computed in non-LTE or in LTE, the temperature law is flat in the outer part of the model. In the constant-temperature regimes, the electron density is such that log N_e < 12.0.

Comparison of Figures 7-4 and 7-5 with Figures 7-1 and 7-2 shows that the parts of the models most important for forming the lines He I 4471 and He II 4541 lie in the parts of the models where log m < -3.0. The temperature is effectively constant here at about 22200 K in the model with T_{eff} = 30000 K and at about 38200 K in the model with T_{eff} = 45000 K.

The continuum near 4500 Å is formed in the layers close to the point where log m = 0.0. Here the electron temperatures are about equal to the effective temperature, and log N_e is near 15.0. The typical temperatures and densities of the regions which are important for forming the selected lines of He I and He II are significantly different from those for forming the continuous spectrum near 4500 Å. The region of formation of lines having smaller f values than those of the selected lines can be estimated by considering the properties of the layers in which t_ν (line center) is even larger than 10^6. Allowing for any reasonable range in f values still produces the result that most of the line profile is formed well outside the place where the continuum is formed. Only for the wings of the He II lines in the model having T_{eff} = 30000 K does the depth of formation for the lines approach that for the continuum.

With main-sequence B stars, the intrinsic $(B\text{-}V)_0$ or $(b\text{-}y)_0$ color indices in the visible range and the size of the Balmer jump are useful criteria for selecting representative model atmospheres. Use of these criteria is not very helpful for O stars, because the measured Balmer jumps are small and they are not sensitively dependent on T_{eff}, while the intrinsic colors of O stars are not well determined in either the *UBV* or the *uvby* photometric system (see Part I). Fitting an observed energy distribution over a long wavelength range between the ultraviolet and the visible spectral range serves to confirm effective temperatures found from integrated fluxes and angular diameters.

This has been demonstrated by Underhill (1982) for a few O stars. It provides probably the most reliable method for confirming a representative model atmosphere for the photosphere of an O star. Selection of a model atmosphere by matching the equivalent widths of the strong lines in the visible range that are used for spectral classification certainly can be done, but the effective temperature found in this way may be systematically higher than the effective temperature corresponding to the integrated flux from the star.

It is sometimes suggested that the atmospheres of O stars are so extended that use of plane parallel layers even when attempting to model only the layers important for forming the continuous spectrum from 1200 Å to 1 μm is unsatisfactory. This matter is discussed further in *Properties of Hot, Static Spherical Model Atmospheres*. The numerical results now available for a few static spherical model atmospheres having the effective temperatures of O stars describe continuous spectra that have shapes which are closely similar to those produced by plane parallel layer models of equivalent effective temperatures. Analysis of the continuous spectrum of O stars from 1200 Å to 1 μm by means of plane parallel layer model atmospheres appears to be a satisfactory procedure (Underhill, 1984b).

C. Analyses of the Absorption Lines of O Stars: Abundances

The method of "coarse analysis" will serve for obtaining a first approximation for the temperatures and densities in an O-type atmosphere. It consists essentially of a curve-of-growth study of the absorption lines in the spectrum. The procedures which are followed and the results which have been attained for O and B stars have been summarized in Chapter XI of *The Early-Type Stars* (Underhill, 1966). The method of "coarse analysis" does not give consistent results for O stars. This may be seen from the results of Oke (1954), who tried some simple modifications in an attempt to improve the internal consistency of his results.

Better consistency between theory and observation is obtained when the methods of "fine analysis" are applied to absorption lines in the visible part of an O-type spectrum. First one selects a representative model atmosphere for the line-forming parts of the atmosphere by the method described above or by finding which (T_{eff}, log g) pair best represents the ionization balance between lines in two stages of ionization of several elements. The spectra usually studied for this purpose are He I/He II, C III/C IV, N II/N III, N III/N IV, O II/O III, and Si III/Si IV. Alternatively, one may match the observed size of the Balmer jump.

After representative values for T_{eff} and log g have been found, the observed equivalent widths of many lines are matched by varying the abundance of the element. In the case that an LTE model atmosphere and an LTE representation for the physics of line formation are used, microturbulence may be introduced in order to make weak lines and strong lines from the same element yield about the same abundance. In the case that non-LTE model atmospheres and non-LTE theory of line formation are used (see, for instance, Auer and Mihalas, 1972), the microturbulent velocity is usually kept equal to zero km s^{-1}.

1. LTE Analyses. The effective temperatures and log g values for the model atmospheres which have been proposed for representing the line-forming regions in the atmospheres of 11 O stars are listed in Table 7-1. The composition of all these model atmospheres is such that N(He)/N(H) is about equal to 0.10. In the model atmospheres of Peterson and Scholz (1971), account is taken of the continuous absorption from the ions of C, N, O, and Ne, while in those of Mihalas (1972b), which are used by Takada (1977), the continuous absorption from the ions of C, N, and O is simulated by a fictitious light element. Takada found that, to obtain consistent results when he was making an LTE analysis of the spectra of α Cam and 19 Cep, he had to introduce a large microturbulent velocity. Underhill and de Groot (1964) and Underhill (1968a) found in their LTE analyses of the spectrum of 10 Lac that they had to introduce a microturbulent velocity of the order of 12 km s^{-1} in order to obtain calculated profiles of lines other than from H I, He I, and He II which looked like the observed ones. They could obtain no match at all for the lines of H I, He I, and He II.

Table 7-1
Effective Temperatures and log g Values for Representative Model Atmospheres

HD	Name	Walborn Spectral Type	T_{eff} (K)	log g	ξ_t (km s^{-1})	Ref.*	T_{eff} (K)	log g	Ref.*
30614	α Cam	O9.5 Ia	29000	3.05	20	1	32500	3.15	1
34078	AE Aur	O9.5 V	37000	4.4	5	2	36000	4.2	3
36861	λ Ori A	O8 III((f))	41000	4.3	5	2	37500	4.0	3
37742	ζ Ori A	O9.7 Ib					31000	3.2	3
47839	15 S Mon	O7 V((f))	43500	4.5	5	2	39000	4.2	3
54662	HR 2694	O6.5 V	45000	4.4	5	2	41000	4.2	3
57682	HR 2806	O9 IV	38500	4.4	5	2	35000	3.9	3
66811	ζ Pup	O4 I(n)f					~50000	4.0	4
164794	9 Sgr	O4: V((f))	54000	4.4	5	2	>50000	4.0	3
209975	19 Cep	O9 Ib	30000	3.30	22.5	1	35000	3.3	1
214680	10 Lac	O9 V	34000	4.0	12	5	37000	4.25	3
214680	10 Lac	O9 V	32000	4.0	12	6			

*References: 1. Takada (1977), 2. Peterson and Scholz (1971), 3. Auer and Mihalas (1972), 4. Baschek and Scholz (1971), 5. Underhill and de Groot (1964), 6. Underhill (1968a).

The effective temperatures listed in Table 7-1 may be compared with effective temperatures deduced from the integrated fluxes and angular diameters of O stars (see Underhill et al., 1979; Underhill, 1982; and Part I of this book). In most cases, the values in Table 7-1 are significantly higher than those found from integrated fluxes.

The LTE analyses showed that a solar composition was about correct for O-type stars. However, there were some discrepancies, by up to about a factor 5. It was not certain whether these discrepancies were due to using an inadequate theory of line formation (i.e., LTE theory) or to real differences in composition from the Sun.

On the whole, however, O-type line profiles predicted by means of LTE model atmospheres and an LTE theory of line formation do not agree well with the observed line profiles. This is particularly evident for the strong lines of H I, He I, He II, Si III, and Si IV which are used to classify type O spectra. The classification criteria for O stars are reviewed in Part I of this book. The predicted variation of the equivalent widths of lines of H I, He I, and He II with T_{eff} and log g in the LTE model atmospheres for O stars can be found from the diagrams and numerical data presented by Auer and Mihalas (1972).

The major areas of misfit with LTE theory are:

1. The Balmer lines of hydrogen persist to type O5 at greater strength than is predicted.

2. The triplet lines of He I are observed to persist at significant strength throughout type O, while the singlet He I lines vanish in the earlier types. The LTE theory predicts that both types of He I lines weaken at rates which are not observed.

3. The Pickering lines of He II are observed to persist at increasing strength from type O9 to at least type O5, although LTE calculations indicate that these lines should reach a maximum around type O6 and then decrease in strength.

4. The Mg II doublet at 4481 A is observed to persist into type O at greater strength than is suggested by LTE calculations using the solar abundance.

5. The lines of Si III at 4552, 4568, and 4574 Å, as well as the lines of Si IV at 4089 and 4116 Å, are observed to be stronger than LTE theory with the solar abundance of silicon predicts relative to weak lines from these spectra.

6. A few lines appear in emission in the spectra of O stars, a fact which cannot be predicted by traditional LTE theory.

2. Non-LTE Analysis. Non-LTE analysis of the line spectrum in the visible range has been carried out for eight O stars by Auer and Mihalas (1972). The values of T_{eff} and log g which they conclude specify model atmospheres which best represent the H I, He I, and He II lines are listed in Table 7-1. The misfits between theory and observation are much reduced when non-LTE model atmospheres and a non-LTE theory of line formation are used (see Auer and Mihalas (1972) for H I, He I, and He II; Kamp (1976, 1978) for Si III and Si IV; and Mihalas (1972c) for Mg II). However, some discrepancies still remain. The amount of disagreement between the theory and observation can best be illustrated by comparing a few observed line profiles in O-type spectra with predicted profiles. Such a comparison is made in the next few paragraphs for the normal, sharp-lined star, 10 Lac. The problem of understanding the significance of the emission lines seen in O-type spectra is treated in later sections of this Chapter.

Spectroscopic information about each line discussed here is given in Table 7-2. The variation with effective temperature of the predicted equivalent widths for these lines is shown in Figures 7-6 and 7-7. In Figure 7-6, the equivalent widths of Auer and Mihalas (1972) are shown for some H I, He I, and He II lines for the case of NLTE/L models having log g = 4.0. In Figure 7-7, the equivalent widths

Table 7-2
Spectroscopic Information for Some Strong Lines in O-Type Spectra

Spectrum	λ(Å)	Transition	J	Lower E. P. (eV)	I.P. (eV)	Ref.[a]
H I (Hα)	6562.796[b]	2 – 3		10.20	13.598	1,2
H I (Hγ)	4340.464[b]	2 – 5		10.20	13.598	1,2
He I	4471.479[b]	$2^3P^o - 4^3D$	2,1 – 3,2	20.96	24.588	3
He I	4471.682	$2^3P^o - 4^3D$	0 – 1	20.96	24.588	3
He I	4921.931	$2^1P^o - 4^1D$	1 – 2	21.22	24.588	3
He I	4387.929	$2^1P^o - 5^1D$	1 – 2	21.22	24.588	3
He II	1640.417[c]	2 – 3		40.81	54.418	2
He II	4685.706[b]	3 – 4		48.37	54.418	2
He II	5411.521[b]	4 – 7		51.02	54.418	2
He II	4541.591[b]	4 – 9		51.02	54.418	2
Si III	4552.616	$4^3S - 4^3P^o$	1 – 2	19.02	33.493	4
Si IV	4088.854	$4^2S - 4^2P^o$	½ – 1½	24.05	45.142	4
Si IV	4116.097	$4^2S - 4^2P^o$	½ – ½	24.05	45.142	4

[a] References: 1. Moore (1972), 2. Garcia and Mack (1965), 3. Martin (1960), 4. Moore (1965).
[b] Mean wavelength in air.
[c] Mean wavelength in vacuum.

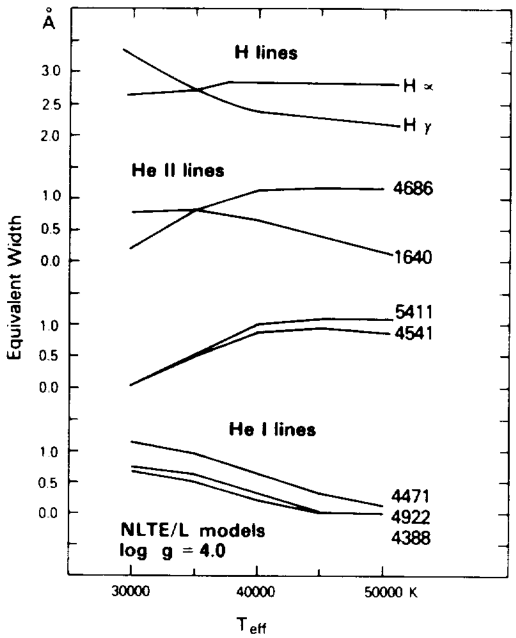


Figure 7-6. The variation with T_{eff} of the equivalent widths of lines of H I, He I, and He II in NLTE/L model atmospheres having log g = 4.0 (from Auer and Mihalas, 1972).

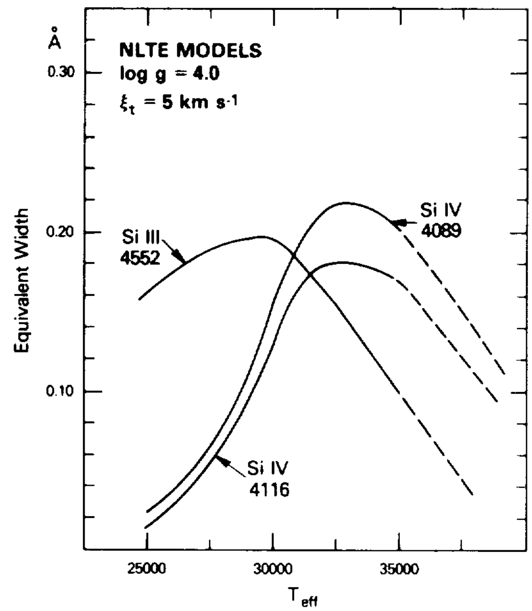


Figure 7-7. The variation with T_{eff} of the equivalent widths of lines of Si III and Si IV in NLTE/L model atmospheres having log g = 4.0. The data (Kamp, 1976) for the case, ξ_t = 5 km s^{-1}, are plotted. The broken lines show extrapolations for $T_{eff} > 35000$.

from Kamp (1976) for selected Si III and Si IV lines are shown for the same models. The case chosen for display is for microturbulent velocity ξ_t, equal to 5 km s^{-1}. An extrapolation of Kamp's results to effective temperatures greater than 35000 K is suggested by means of broken lines.

Although the equivalent width of Hγ decreases as the effective temperature increases, the equivalent width of Hα reaches a maximum value near 2.83 Å at $T_{eff} \approx 37000$ K, and it remains at this value. The line He II 1640 reaches maximum equivalent width at 35000 K and then decreases in strength. The line He II 4686, however, reaches maximum strength at about 40000 K and remains strong. The lines of the Pickering series of He II at 5411 and 4541 Å increase in strength as the effective temperature increases, reaching a plateau of strength between 45000 and 50000 K. The selected lines of He I decrease in strength as the effective temperature increases, the singlets effectively disappearing by 45000 K, although the triplet line at 4471 Å remains significant with an equivalent width $\geqslant 0.15$ Å. When a larger value of log g is used, the equivalent widths of these lines increase, since the lines of H I, He I, and He II are sensitive to Stark effect. The line wings are deepened, but the line center is changed very little.

The Si III line at 4552 Å reaches maximum equivalent width at about 30000 K; it should become difficult to detect when $T_{eff} > 37500$ K. The Si IV lines at 4089 and 4116 Å reach maximum equivalent width when the effective temperature is about 33000 K; they may be expected to become difficult to detect when $T_{eff} > 45000$ K.

In Figures 7-8 to 7-13, observed profiles of the lines in Table 7-2 from the spectrum of 10 Lac (de Groot and Underhill, 1964; Underhill, 1968a) are compared with the predicted line profiles (Auer and Mihalas, 1972; Kamp, 1976) from NLTE/L models having log g = 4.0,

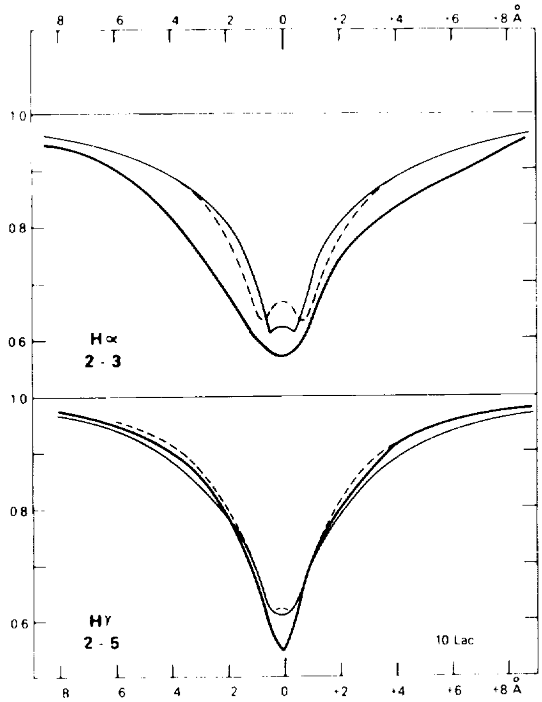


Figure 7-8. Profiles of Hα and Hγ in the spectrum of 10 Lac compared with the predicted profiles for models (35000, 4.0, NLTE/L) and (40000, 4.0, NLTE/L). The observed profiles (Underhill, 1968a) are shown by a heavy line, the predicted profiles for T_{eff} = 35000 K by a thin line, and the predicted profiles for T_{eff} = 40000 K by a broken line.

$T_{\rm eff}$ = 35000 and 40000 K. The observed line profiles are from 4.5 Å mm^{-1} spectrograms obtained at the Mount Palomar Observatory, and they are believed to be reliable. (See de Groot and Underhill, 1964; and Underhill, 1968a for a critical discussion of these data.) No significant systematic error in the photometry is suspected. The observed profile of He II 1640 has been found from International Ultraviolet Explorer (IUE) image SWP 6286.

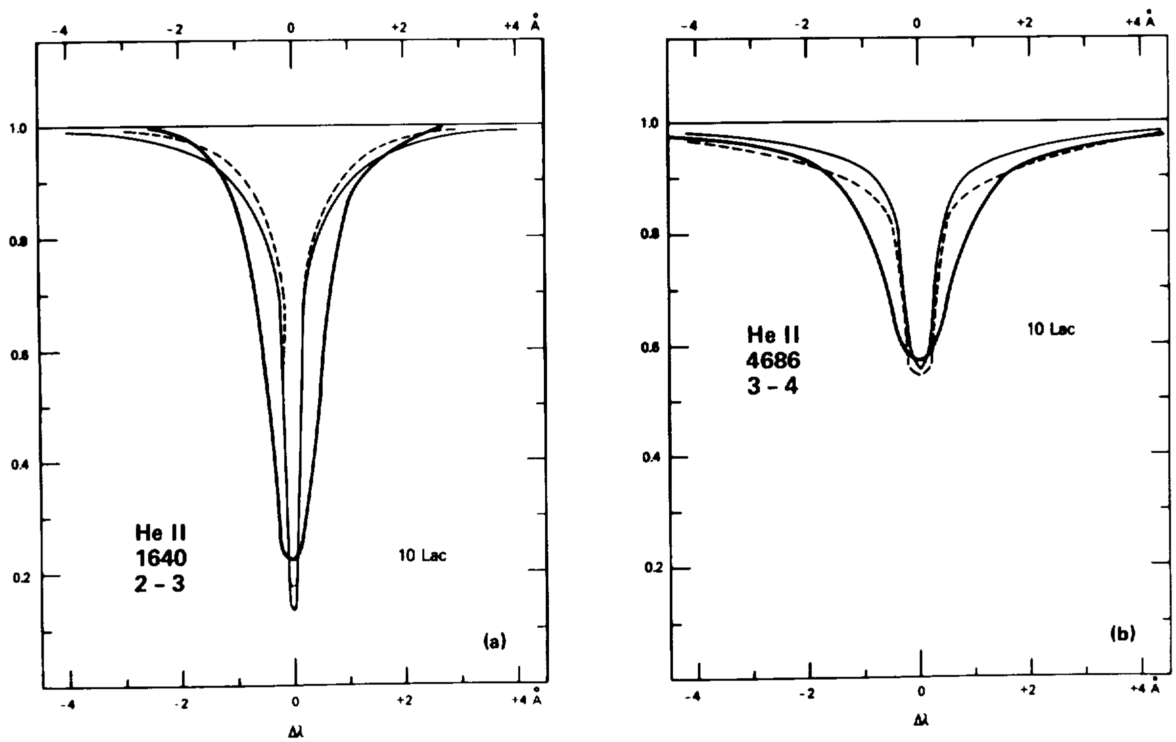


Figure 7-9. Profiles of He II 1640 and He II 4686 in the spectrum of 10 Lac compared with the predicted profiles for T_{eff} = 35000 and 40000 K. The key to the meaning of the different lines is as in Figure 7-8. The profile for He II 1640 is from IUE image SWP 6286.

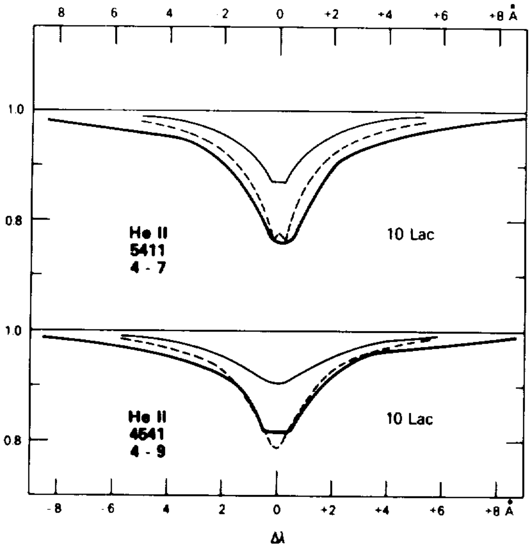


Figure 7-10. The observed profiles of two lines of He II from the Pickering series in the spectrum of 10 Lac compared to computed profiles. The meaning of the lines is as in Figure 7-8.

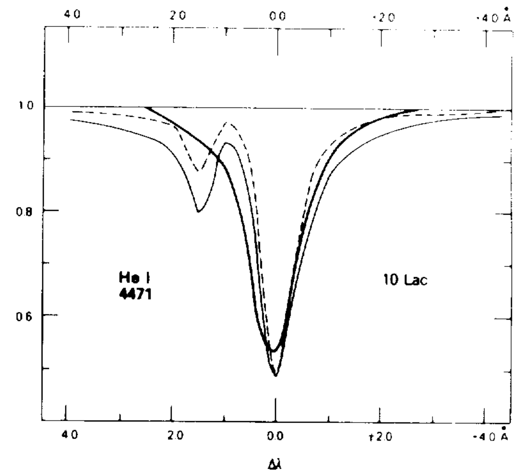


Figure 7-11. The observed profile of the triplet He I line at 4471 Å in the spectrum of 10 Lac compared with predicted profiles. The meaning of the lines is as in Figure 7-8.

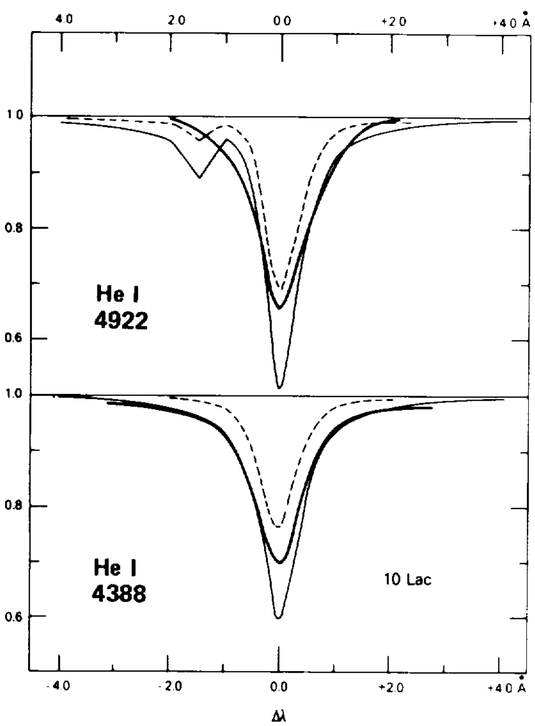


Figure 7-12. The observed profiles of two singlet He I lines in the spectrum of 10 Lac compared with predicted profiles. The meaning of the lines is as in Figure 7-8.

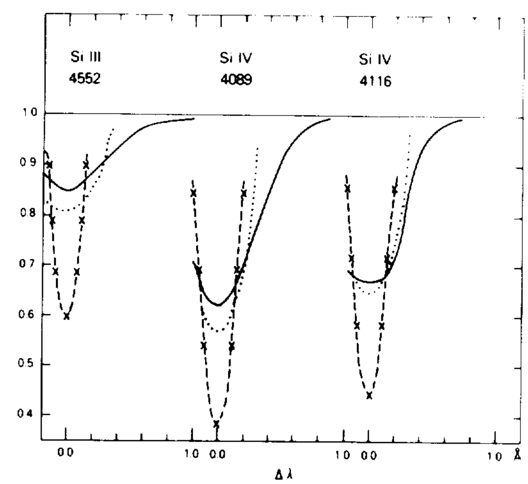


Figure 7-13. Observed profiles of Si III 4552, Si IV 4089, and Si IV 4116 in the spectrum of 10 Lac compared with predicted line profiles for model (35000, 4.0, NLTE/L). The heavy line shows the observed profiles (de Groot and Underhill, 1964). The broken line shows the part of the profile listed by Kamp (1976), and the dotted line shows this profile broadened by a rotational velocity of 23 km s^{-1} (Uesugi and Fukuda, 1970) according to the simple broadening formula of Unsöld (1955). The limb-darkening coefficient, β, was put equal to 0.5.

The effective temperature of 10 Lac, found from the integrated flux and an estimated angular diameter, is 35400 K according to Underhill et al. (1979), and the shape of the spectrum from 1300 Å to 1.25 μm agrees well with the shape of the continuous spectrum from the Kurucz (1979) model atmosphere having T_{eff} = 35000 K, log g = 4.0 (Underhill, 1984b). Precise representative values of T_{eff} and log g for a model atmosphere for 10 Lac cannot be selected by reproducing the shape of the spectrum over the observed wavelength range from the ultraviolet to the near infrared. This is because the uncertainties in the observed energies are of the order of $\pm$10 percent, and the uncertainties in our knowledge of the wavelength-dependent interstellar extinction from the ultraviolet to the near infrared are of the same order of magnitude, while the predicted shape of the spectrum is not sensitively dependent on T_{eff} and log g.

Intercomparison of the predicted continuous spectra from the models of Kurucz (1979), which have T_{eff} = 35000 K, log g = 4.0 or 4.5, and T_{eff} = 40000 K, log g = 4.0 or 4.5, shows that over most of the observed spectral range the *shape* of the spectrum is very similar. In particular, continua from the models (35000, 4.5) and (40000, 4.0) have almost identical shapes. These continua differ from the continuum from model (35000, 4.0) by becoming about 0.14 mag brighter at 1250 Å when all three models are matched at 5556 Å. The shapes of the continua from the NLTE/L model atmospheres of Mihalas (1972b) having the same values of T_{eff} and log g are related to each other in about the same way. Consequently, since the uncertainties in the observed energies are of the order of $\pm$10 percent, and since our knowledge of the interstellar extinction law is particularly uncertain shortward of 1500 Å, the observed shape of the continuum for 10 Lac cannot be used to differentiate between the Mihalas (1972b) NLTE/L models and the Kurucz (1979) models which have T_{eff} = 35000 or 40000 K and log g = 4.0 or 4.5. All eight models represent well the observed shape of the continuous spectrum of 10 Lac.

To show how well profiles predicted using the non-LTE theory of line formation and non-LTE model atmospheres compare with observed profiles, we shall compare some observed line profiles given by Auer and Mihalas (1972) and by Kamp (1976). We compare with the profiles from models having log g = 4.0, T_{eff} = 35000 and 40000 K. We have chosen to use model atmospheres with log g = 4.0 because Conti and Leep (1974) have classified 10 Lac as O8 III. We understand this to mean that, in their opinion, the spectrum of 10 Lac suggests a luminosity somewhat above the lowest values found at type O8. The spectral type assigned by Walborn (1971) is O9 V. The star, 10 Lac, has a projected rotational velocity of 23 km s^{-1} (Uesugi and Fukuda, 1970). We ignore rotational broadening for the Stark broadened lines of H I, He I, and He II. If the theoretical line profiles were broadened to take account of rotation, their central intensities would be made a little shallower, but the shape of the wings more than 0.5 Å from the line center would be unchanged.

The observed profiles for Hα and Hγ in the spectrum of 10 Lac (Underhill, 1968a) are shown in Figure 7-8 together with theoretical profiles from Auer and Mihalas (1972). The observed Hα line is broader and deeper than either computed line. Both of the computed profiles show rises in their centers which occur because the outer layers of the selected model atmospheres are optically very thick in wavelengths near the center of Hα. The observed profile for Hγ agrees well with the predicted profiles at distances of more than 1 Å from the line center. This suggests that the selected value of log g is representative. The predicted wings for Hγ are not very sensitive to T_{eff}. Auer and Mihalas (1972) have published a comparison between the observed profile of Hβ and predicted profiles for models with T_{eff} = 35000 and 40000 K. The observed Hβ profile is deeper than the predicted profiles at all wavelengths, but the discrepancy is not so great as at Hα. These misfits for the Balmer lines appear to be real; there is no reason to suspect that the observed profiles contain systematic errors of a type and magnitude which would produce the observed discrepancies.

The observed profiles in the spectrum of 10 Lac for He II 1640 and He II 4686 are compared with predicted profiles in Figures 7-9a and 7-9b. Discrepancies similar to those found for Hα and Hβ are evident. Almost all of each observed profile is deeper than the predicted profiles. It was noted above that broadening the theoretical profiles by a small amount of rotation would bring the centers of the predicted profiles into better agreement with the centers of the observed lines. In any case, all of the observed line centers are contaminated by the presence of a small amount of scattered light. Some 2 or 3 percent of the neighboring continuum is typical for grating spectrographs of the sort used to make the observations. It seems clear that the observed profiles for He II 1640 and He II 4686 have significantly broader cores than those predicted by non-LTE theory used with static non-LTE model atmospheres.

The observed profiles in the spectrum of 10 Lac for He II 5411 and He II 4541, lines of the Pickering series, are compared with predicted profiles in Figure 7-10. The line at 5411 Å is observed to be stronger and deeper than either of the predicted profiles. The predicted absorption in both He II lines is a sensitive function of the adopted effective temperature. The centers of these lines are formed in the outermost layers of the selected model atmospheres. The observed profile for He II 4541 agrees rather well with the predicted profile from the model with T_{eff} = 40000 K. However, the wings of both these He II lines are observed to be deeper than the predicted wings at all distances from the line center.

The profile of the triplet He I line at 4471 A is compared with predicted profiles in Figure 7-11; observed and predicted profiles of the singlet He I lines at 4922 and 4388 Å are compared in Figure 7-12. The theory of the Stark effect of He I used by Auer and Mihalas (1972) is not entirely satisfactory for representing the forbidden components, $2^{1,3}P^o$ - $4^{1,3}F$, at 4469 and 4920 Å. An improved theory has been used by Mihalas et al. (1974) to predict profiles which are in better agreement with the observations. However, Mihalas et al. did not publish predicted profiles for the He I lines at T_{eff} > 27500 K. Consequently, the results of the improved theory cannot be compared with the spectrum of 10 Lac.

Let us direct our attention chiefly to the agreement between theory and observation on the longward side of the He I lines. Here the Stark broadening theory used by Auer and Mihalas should be reliable. The observed profile for He I 4471 lies between the profiles predicted for 35000 and 40000 K. Auer and Mihalas (1972) noted that the observed equivalent width of the 2^3P^o - 3^3D He I line at 5876 Å is considerably larger than the predicted values for a likely effective temperature. The equivalent width of the profile shown by Underhill (1968a) is 1.043 Å, while the equivalent widths predicted by Auer and Mihalas (1972) for NLTE/L models with log g = 4.0 and T_{eff} = 35000 and 40000 K are 0.662 and 0.610 Å, respectively.

The observed profiles for the singlet He I lines at 4922 and 4388 Å are more like the profiles for T_{eff} = 35000 K than those for 40000 K. In the case of the singlet He I lines, the predicted profile is rather sensitive to the adopted value of T_{eff}. The observed equivalent width of the profile shown by Underhill (1968a) for the 2^1P^o - 3^1D line of He I at 6678 Å is 0.768 Å, while the predicted equivalent widths for this line are 0.647 and 0.402 Å, respectively, for T_{eff} equal to 35000 and 40000 K.

The observed line profiles for He I 4471, 4922, and 4388 in the spectrum of 10 Lac suggest that a model having an effective temperature a little larger than 35000 K and log $g \approx 4.0$ would produce good agreement with the profiles observed for these lines of He I in the spectrum of 10 Lac. However, the observed equivalent widths of the *leading* members of the singlet and the triplet diffuse series of He I are much larger than the values which are predicted. The discrepancy is significantly greater than any uncertainty in the measured equivalent widths. The observed profiles for He I 5876 and He I 6678 cannot be compared with the predicted profiles because Auer and Mihalas did not publish predicted profiles for these lines.

In Figure 7-13, the profiles of Si III 4552, Si IV 4089, and Si IV 4116 observed in the spectrum of 10 Lac (de Groot and Underhill, 1964) are compared with profiles predicted by Kamp (1976) for the case T_{eff} = 35000 K, log g = 4.0, and ξ_t = 5 km s^{-1}. The symmetrical profile corrected for instrumental broadening has been reconstructed from the information given by de Groot and Underhill. The profiles for the cores of the predicted lines are listed by Kamp. It is clear that these cores (shown by broken lines in Figure 7-13) are much narrower than the cores of the observed profiles. An impression of how much broadening will be caused by a rotational velocity of 23 km s^{-1} is shown by the dotted lines. The dotted lines were obtained by calculating the effects of rotational broadening by means of the simple formula given by Unsöld (1955). The equivalent width of each rotationally broadened profile has been made equal to the equivalent width for the line given by Kamp. Because rotational broadening by 23 km s^{-1} produces a profile with a full width at half maximum (FWHM) which is comparable to the FWHM of the unbroadened theoretical profiles, the rotationally broadened profiles should be found by a more correct theory of line broadening than that used here—one such as the formulation used by Stoeckley and Mihalas (1973). Such a theory would give profiles with less steep edges than those shown in Figure 7-13. Nevertheless, it is clear that the broadening caused by a rotational velocity close to 23 km s^{-1}, which is the value listed for 10 Lac by Uesugi and Fukuda (1970), will not produce wings such as those observed for Si III 4552, Si IV 4089, and Si IV 4116 in the spectrum for 10 Lac.

The selected lines of Si III and Si IV are stronger than the NLTE theory predicts, their equivalent widths being 0.101, 0.309, and 0.229 Å, respectively (Underhill and de Groot, 1965), while the equivalent widths (Kamp, 1976) of the theoretical lines shown in Figure 7-13 are 0.0706, 0.1775, and 0.1450 Å, respectively. Using a model atmosphere with a higher effective temperature would make the discrepancy even greater.

In summary, the strong absorption lines in the spectrum of 10 Lac, some of which are used for determining the spectral type of the star (see Part I) are not generally well represented by non-LTE line-formation theory using model atmospheres which represent the shape of the continuous spectrum reasonably well. The discrepancies would not be eliminated by going to T_{eff} = 37000 K and log g = 4.2 as Auer and Mihalas (1972) suggest should be done. The largest discrepancies are seen in the leading members of series of lines from abundant species. Typically, the observed profiles of leading members in the spectrum of 10 Lac are deeper and wider than theory predicts. This has been demonstrated here for Hα, He II 1640, and He II 4686; it may be inferred for He I 5876 and He I 6678.

Lines in the visible range from relatively low-lying levels in other abundant species such as Si^{+2} and Si^{+3}, as well as the higher members of the Pickering series of He II, are also broader than theory suggests. However, the profiles of other lines such as He I 4471, 4922, and 4388 are predicted relatively well.

Our theoretical understanding of the observed absorption lines in normal O-type stars of the main-sequence band is far from satisfactory. We seem to be seeing part of the mantle by means of the lines listed in Table 7-2. However, little detailed theory exists at this time for modeling mantles and for studying line formation in them. Further information on this point can be found in Underhill (1984b).

Analysis of the spectra of subluminous O stars can be carried out in exactly the same manner as that for normal Population I O stars. However, the detail in which the analysis can be done is limited by the relatively poor spectral resolution of the available spectra of subluminous O stars. Dufton (1972) demonstrated that one must use non-LTE physics when studying the lines of H, He I, and He II in the spectra of sdO stars, and he showed that helium seemed to be overabundant in the spectrum of the subdwarf HD 49798.

A systematic study of the spectra of subluminous O stars using non-LTE physics has been carried out by the group at Kiel University. Because most of their spectra are at 29 Å mm^{-1} with a spectral resolution of the order of 1.6 Å, this group has concentrated on the analysis of the blended H and He II lines at Hβ and Hγ, He I 4471 and 4713, and He II 4686. They have found values for log T_{eff} and log g for each star, as well as the quantity, $y = N(\mathrm{He})/(N(\mathrm{H})+N(\mathrm{He}))$. The spectra of 15 subluminous O stars (Kudritzki and Simon, 1978; Simon, 1979; Kudritzki et al., 1980; Baschek et al., 1980; Hunger et al., 1981) and six central stars of planetary nebulae (Méndez et al., 1981) have been analyzed. In the case of the subluminous O stars, the effective temperatures fall in the range from 35000 to 55000 K, while log g falls in the range from 4.5 to 6.3; for the O-type central stars, T_{eff} is higher, lying in the range from 50000 to 75000 K, while log g lies in the range from 4.5 to 7.0.

So far as fractional helium content, y, is concerned, there is a wide spread. Four of the sdO stars and three of the central stars have y of the order of 0.03; one sdO star and three central stars appear to have normal helium content, y of the order of 0.1; eight of the sdO stars have y in the range from 0.2 to 0.7, while two have $y = 1.0$. The authors suggest that apparent helium overabundance may mean that the star has lost most of its original hydrogen-rich atmosphere, while helium underabundance may indicate that gravitational settling of helium has occurred in the atmosphere of the star.

Little analysis of spectral lines from elements other than H and He has been done. Gruschinske et al. (1980) have studied the strength of the C IV resonance lines in the spectra of three sdO stars (HD 49798, HD 127493, and BD+75°375). They find that carbon appears to be deficient relative to solar abundances by a factor of 10. In addition, Simon et al. (1980) have determined by a non-LTE analysis of lines of the doublet system of N III and of the Si IV resonance lines that nitrogen is strongly overabundant, while silicon appears to have solar abundance in the atmosphere of these stars.

D. Discrepant Lines in O-Type Spectra

Two types of discrepancies occur between the equivalent widths and profiles of the chief lines observed in O-type spectra and the equivalent widths and profiles predicted for these lines by means of traditional theory (LTE or non-LTE) using model atmospheres composed of plane parallel layers of gas in hydrostatic and radiative equilibrium. The first is the exceptional strength in absorption and/or in emission of the leading members of series of lines in spectra formed by one or two electrons outside a charged nucleus or outside a core composed of a nucleus and one, two, or three shells of tightly bound electrons corresponding to the arrangements of electrons denoted by $1s^2$, $2s^2$, and $2p^6$. Typical lines of this group come from the H I, He I, He II, C IV, N V, O VI, and Si IV spectra. The second is the occurrence in more complex spectra such as C III and N III of a few selected lines in emission. Emission is not seen in lines which follow each other in the cascade chains identified by the emission lines which are seen. Typical lines of the second group are often seen in Of spectra. They are C III 5696, C III 9701-9719, and N III 4634, 4641, and 4642.

The first type of discrepancy—the leading-member discrepancy, whether in absorption or in emission, or both—can be understood qualitatively in terms of an extended atmosphere which is opaque in frequencies near the line center, but which is transparent in continuum frequencies in the neighborhood of the line. To understand the second type of discrepancy—the single-line discrepancy—one must study the individual radiative and collisional processes which occur in a stellar atmosphere and determine how they influence the distribution of the atoms/ions over their several possible energy states.

The theory of line formation in an expanding sphere of gas around a photosphere is reviewed in *The Line Spectrum from Moving Three-Dimensional Model Atmospheres* later in this chapter. It provides an explanation for "wind" profiles seen in O-type spectra for the resonance lines from abundant species and for the emission seen at most lines which show discrepancies of the "leading-member" type. A wind profile consists of a shortward displaced absorption trough accompanied by an emission peak on the longward side of the line. A line showing such a profile is often called a P-Cygni-type line.

To understand the origin of a broad undisplaced emission line at Hα with a weak, or absent, absorption component, one must postulate an extended atmosphere with large velocities directed toward and away from the observer, as well as a small population of hydrogen atoms in the $n = 2$ level. How the postulated structure of gas and its associated velocity field and electron temperatures has come into existence is a separate, and usually unanswered, question.

Insight into whether a line will appear in emission or not may be obtained by considering the value of the source function for the line along all lines of sight between the observer and parts of the object sending radiation in the line frequencies toward the observer. In terms of a two-level atom, one may readily show that, in each element of volume, the source function for a line is proportional to n_{upper}/n_{lower}. Here n_{upper} is the population by number of the atom or ion in the upper level of the line being considered, and n_{lower} is the population of atoms or ions in the lower level of the line. If n_{upper}/n_{lower} is large along most lines of sight which bring radiation in the specified line to the observer, the line will appear in emission. If n_{upper}/n_{lower} is small along most lines of sight, then the line will be seen in absorption. Emission appears in some lines simply because the atmosphere is detectable over a much larger volume in line frequencies than it is in continuum frequencies—the planetary-nebula picture. In such a case, the profile of the line may be determined more by the velocity field in the extended atmosphere than by any physical cause of line broadening such as Stark effect or thermal Doppler broadening. The intensity of light is stronger in line frequencies than in the nearby continuum frequencies because there are many more lines of sight per element of area on the sky bringing light from the line-emitting region to the observer than there are lines of sight bringing continuum radiation at a significant level of intensity.

On the other hand, the same geometrical configuration may be sending radiation in line frequencies and in continuum frequencies to the observer. This is the case for a configuration of plane parallel layers of gas. Even then, emission may occur in some lines if, owing to particular cascade and excitation chains, some levels are overpopulated relative to others. In one line, n_{upper}/n_{lower} may be large, while in another line in the same spectrum, this ratio may be small. In fact, depending on the electron density and the local radiation field, a selected line may appear in emission in some model atmospheres but not in others. What happens depends sensitively on the relative efficiencies of the recombination, excitation, and ionization processes which depend on the local radiation field and the electron density. This is what happens in "single-line" discrepancies.

Mihalas and Hummer (1973) have investigated the detailed balancing of the radiative and collisional processes in the spectrum of N III in O-type model atmospheres. They have found that the 3^2P^o - 3^2D lines of N III (multiplet 2) at 4634, 4641, and 4642 Å will appear in emission in the spectra of hot model atmospheres with low values of log g, while the 3^2S - 3^2P^o lines at 4097 and 4103 Å (multiplet 1) remain in absorption, exactly as is observed in some O stars. The key factors in producing relative populations of the levels of N III such that multiplet 2 appears in emission while multiplet 1 appears in absorption are: (1) dielectronic recombination from the low-lying

$2s2p(^1P^o)$ 3d autoionizing states, with stabilizing cascades via 3d → 3p (multiplet 2), and (2) draining of the population in the 3p state by two-electron jumps which couple the 3p state to the $2s2p^2(^2S, ^2P, ^2D)$ states and thus prevent emission in the 3s - 3p lines (multiplet 1). In all of the model atmospheres used by Mihalas and Hummer, the He II line at 4686 Å is strongly in absorption (Auer and Mihalas, 1972). Consequently, one may infer that the mechanism investigated by Mihalas and Hummer serves to explain the difference between the normal O stars and the O((f)) stars. In O(f) stars, the He II line at 4686 Å is neutralized by emission, while in Of stars, the He II line is firmly in emission. To neutralize the absorption in He II 4686 and to obtain He II 4686 in emission, one must postulate the presence of an extended atmosphere, as discussed above. A temperature rise may also be postulated. No calculations have been done to show how such a geometrical configuration would affect the intensities of the key lines of N III which are in the cascade chains which are initiated by dielectronic recombination.

The spectrum of C III in O stars shows several characteristic features which are similar to the features which we have been discussing in N III. In particular, the C III lines at 5696 and 9701 to 9719 Å are seen in emission in Of stars, but other C III lines appear only in absorption. Cardona (1978) has carried out a non-LTE study of the singlet C III spectrum in O-type model atmospheres, and he has shown that the combination of emission in the 3^1P^o - 3^1D line at 5696 Å (multiplet 2) and absorption in the 3^1S - 3^1P^o line at 8500 Å (multiplet 1.01) can occur in the theoretical spectra from the NLTE model atmospheres of Mihalas (1972b) having T_{eff} in the range from 30000 to 50000 K and log g in the range 3.0 to 4.5. The key factor in keeping the line at 8500 Å in absorption is the draining of the 3^1P^o level by two-electron jumps to the $2p^2(^1S, ^1D)$ states. This prevents the 3^1S - 3^1P^o line from going into emission. Dielectronic recombination is an important factor in generating the cascades through the 3^1P^o - 3^1D line only for the hottest model atmospheres. In model atmospheres having T_{eff} near 30000 K, simple recombination coupled with the two electron jumps will result in emission in multiplet 2 but not in multiplet 1.01.

The work of Mihalas and Hummer (1973) on the N III spectrum and of Cardona (1978) on the C III spectrum illustrates that the particle and radiation densities seem to be such in the line-forming regions of normal O stars that *particular* radiative and collisional processes are important for causing what we see. This is especially so for lines which have been empirically selected for classifying O-type spectra. The work of Auer and Mihalas (1972) demonstrated the same thing for the chief lines of H I, He I, and He II.

We have shown that even in a normal O star like 10 Lac, the available theoretical predictions are inadequate to account for what is observed. In the strong lines used for spectral classification, we are seeing the effects of line formation in the *mantles* of O stars. In the case of O(f), Of, Of$^+$ and Of*stars (see Part I for the definition of these types), this is even more the case than it is for normal O stars. The theoretical studies using non-LTE physics and non-LTE model atmospheres demonstrate that the typical features by which O-type spectra are classified are formed predominantly in the mantle of the star. Linking an O star to a traditional model atmosphere by means of predicting the relative intensities of one or two conspicuous lines correctly is a procedure which is unlikely to yield reliable values for T_{eff} and log g (i.e., values which can safely be used to infer the evolutionary state of the star). The strengths of the most prominent features in the spectra of O stars are determined by conditions in the mantle of the star. Until we understand how conditions in the mantle of a star are related to the evolutionary state of the star, the spectral type of an O star cannot be used to imply accurately the conditions in the photosphere of the star. In particular, it seems hazardous, in the extreme, to couple a single mass, or even a limited range of mass, with any one spectral type in the O range. The O stars are like the

Wolf-Rayet stars in the following important respect. The most striking features in their spectra are formed in their mantles, not in their photospheres. Most of the available high-temperature model atmospheres are of little relevance for interpreting the *lines* by which Population I O and Wolf-Rayet stars are classified. Additional information on this point can be found in Underhill (1984b) and in Bhatia and Underhill (1986).

E. C III and N III Emission as a Function of Effective Temperature

In Figure 7-14, the predicted equivalent widths of the lines N III 4634 and C III 5696 are shown as functions of the effective temperature, together with the ratio of the equivalent widths of the resonance multiplets of N III and C III. These data have been taken from the work of Mihalas and Hummer (1973) and of Cardona (1978). They were calculated using the non-LTE model atmospheres with log g = 4.0 of Mihalas (1972b). A positive equivalent width means that the line is in absorption; a negative value means that the line is in emission. The resonance multiplets—three blended lines close to 990 A in the case of N III and one line at 977 A in the case of C III—are in absorption in all the model atmospheres. Since the resonance lines are fairly strong in all the model atmospheres investigated, they are probably on the square-root part of the curve of growth or near the bend from the flat part to the square-root part. We shall use the ratio of equivalent widths, W(990)/W(977), as an index of the relative populations of N^{+2} and C^{+2} ions. This index varies about as $[N(N^{+2})/N(C^{+2})]^{1/2}$.

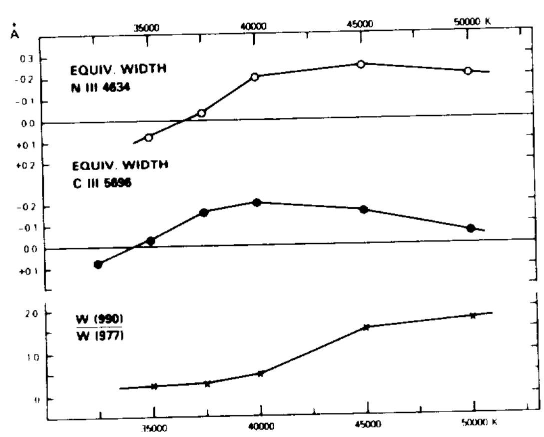


Figure 7-14. The strengths of lines of N III and C III in NLTE plane parallel layer model atmospheres as functions of effective temperature. Log g = 4.0 for all the model atmospheres used to obtain these data (from Mihalas and Hummer, 1973; and Cardona, 1978).

We see from Figure 7-14 that the N III lines which are used to assign the "f" modification to an O spectral type reach maximum strength when the plane parallel layer model atmospheres have T_{eff} in the range from 40000 to 50000 K. Emission in the C III line at 5696 Å has maximum intensity in model atmospheres having T_{eff} in the range 37500 to 45000 K. This is consistent with the observations by Underhill (1955), which indicate that the C III line is not detected in early O spectral types. From the bottom panel of Figure 7-14, we infer that the number of N^{+2} ions in the atmosphere increases rapidly relative to the number of C^{+2} ions when the effective temperature exceeds 40000 K.

The *lines* of the N III and C III spectra are formed in layers of the model atmospheres which are outside the layers in which the continuous spectrum is formed. From our experience with the data displayed in Figures 7-1, 7-2, 7-4, and 7-5, we expect that the most important layers of the models for forming the N III and C III lines under discussion are those in which the electron temperature is nearly constant. Assuming this to be so, we find the tabulations of Mihalas (1972b) that emission in the N III multiplet at 4634, 4641, and 4642 Å is strongest for electron temperatures in the range from 34000 to 41000 K, while the C III line emission is strongest when the electron temperatures are in the range from 30500 to 37000 K. In the outer constant-temperature parts of the non-LTE model atmospheres of Mihalas having log g = 4.0, the electron densities range from about 10^9 to $10^{12} cm^{-3}$.

The variation of the equivalent widths of the resonance lines of N III and C III also points to the same conclusion concerning the temperature ranges favoring N III and C III. The C III spectrum dominates at relatively low electron temperatures near 32000 K. When the electron temperature in the gas exceeds about 37000 K, N III dominates. As the electron temperatures in the gas increase further, we expect lines of C IV and N IV to become strong. Eventually, at electron temperatures greater than 50000 K, the dominant spectra are expected to be N IV, N V, and C IV, with C III absent or undetectably weak. In fact, the calculations of Cardona (1978) show that when the electron temperatures in the model atmosphere are of the order of 43000 K, the lines of C III have negligible strength.

The model atmospheres which Cardona (1978) and Mihalas and Hummer (1973) used are constructed of plane parallel layers of gas of normal composition. Going to a spherical model atmosphere of large volume and projected surface area relative to the photosphere of the star may cause all of the lines to appear in emission. However, the relative numbers of nitrogen and carbon ions in each stage of ionization will remain about the same as those calculated for the plane parallel layers, so long as the electron density remains in the range 10^{10} to $10^{12}cm^{-3}$ and the composition is normal in the large extended atmosphere.

This conclusion is relevant for understanding the meaning of the differences between WN and WC spectra. It suggests that there will be very few ions present which can radiate the C III spectrum when the electron temperatures in the atmosphere are greater than about 45000 K and the electron density is in the neighborhood of 10^{10} to 10^{12}. Many ions radiating the N III, N IV, N V, and C IV spectra are to be expected to be present, however. At very high electron temperatures and moderate densities, the spectrum of N III will disappear and that of C IV will weaken as most of the carbon becomes C^{+4} and C^{+5} ions. The presence of C^{+4} ions may be inferred when one sees a recombination spectrum of C IV; the presence of C^{+5} ions might be inferred if one saw emission in the 2^3S - 2^3P^o multiplet at 2271, 2277, and 2278 Å and in the 2^1S - 2^1P^o line at 3527 Å.

F. Conclusions

From the work reviewed in the preceding paragraphs, it is clear that the *shapes* of the predicted continuous spectra from traditional model atmospheres having normal (solar) composition match well the shapes observed for the continuous spectra of O stars over the range from 1200 Å to 1 μm. In addition, the *amount* of energy predicted in the visible range by traditional model atmospheres, when compared with the amount of energy that is observed, is consistent with what is known about the angular diameters of O stars from direct measurement. Because the predicted continuous spectra from LTE and non-LTE model atmospheres having the same T_{eff} and log g have very nearly the same shape and intensity over the range from 1200 Å to 1 μm, it is adequate to use the predictions from an LTE model atmosphere when one desires to interpret the continuous spectrum from an O star. This conclusion has practical importance because the detail published by Kurucz (1979) about the continuous spectra of his LTE line-blanketed model atmospheres for O stars is much greater than that published by Mihalas (1972b) for non-LTE model atmospheres having the same values of T_{eff} and log g.

When one desires to interpret the equivalent widths and profiles of the absorption lines in O-type spectra, it is advisable to use non-LTE model atmospheres and a non-LTE theory of line formation because the lines are formed in the outer parts of the atmosphere in which particular radiative processes can be at least as important as particular collisional processes for establishing the distribution of the atoms and ions over their sets of energy levels. No comprehensive set of non-LTE analyses of O-type spectra for abundances has been published, but all indications are that the composition of Population I O stars is essentially normal. The non-LTE analyses by Takada (1977) of the spectra of two late O-type supergiants point in this direction.

Although the use of a non-LTE theory of line formation reduces the discrepancies between the strengths and shapes of the absorption lines observed in O-type spectra and the values predicted by theory, it does not remove all of them. The chief discrepancy remaining for absorption lines is that the leading members of H I, He I, and He II series of lines are observed to be deeper and wider in the spectra of normal O stars of the main-sequence band than they are predicted to be. In other O stars, these same leading lines are sometimes seen to be rather strongly in emission; other lines such as C III 5696, N III 4634, 4641, and 4642, N IV 4058, Si IV 4089, and Si IV 4116, and the unidentified lines at 4485.7 and 4503.7 Å also appear in emission in some stars.

Traditional model atmospheres having a geometrical configuration of plane parallel layers cannot produce significant emission at the leading members of the H I, He I, and He II series. It has been argued above that the discrepancies in the leading members, whether in absorption or in emission, as well as the appearance of other lines rather strongly in emission, are due to these intrinsically strong lines from abundant atoms and ions being formed mainly in the mantle (outer atmosphere of the star), and that the physical state of the mantle is not well represented by the traditional model atmospheres which are selected to represent parts of the spectrum formed in the photosphere, such as the continuous spectrum. Although use of plane parallel layer non-LTE model atmospheres and a non-LTE theory of line formation will permit certain lines to appear weakly in emission as the result of particular chains of radiative and collisional processes, such effects cannot account for strong emission in many lines in the spectrum, as seen, for instance, in Wolf-Rayet spectra, nor can it account for strong, broad emission in the leading members of the series of H I, He I, and He II lines.

To account for strong emission lines, one must postulate the presence of a mantle (an outer part of the atmosphere) which is larger when viewed in line frequencies than when viewed in continuum frequencies. If the electron temperature in the mantle is high, lines from ions requiring much energy for their formation may be expected to be strong. Emission in such lines is believed to be chiefly due to stabilizing cascades following the recombination of an ion and an electron into a high level of the next lower ion.

The fact that some O and Wolf-Rayet stars show a detectable amount of excess infrared radiation and radio radiation (see below) also requires one to postulate the presence of a mantle or extended outer atmosphere. The amount of radiation that is measured at very long wavelengths is many times in excess of the amount which is predicted by a model atmosphere which predicts accurately the amount of continuum radiation which is observed from 1200 Å to about 1 μm.

Consequently, it has become apparent that if the full spectrum of O and Wolf-Rayet stars is to be understood, it is necessary to investigate what type of a spectrum may be expected from an extended model atmosphere surrounding a model photosphere which radiates a spectrum like the continuous spectrum observed for an O or a Wolf-Rayet star. In the next section, we shall present the chief theoretical results on the continuous spectra from spherical model atmospheres; after that, we present the chief theoretical results concerning the equivalent widths and profiles of lines predicted using spherical model atmospheres.

III. THE CONTINUOUS SPECTRUM FROM SPHERICAL MODEL ATMOSPHERES

A. Introduction

The earliest investigations of the differences in the spectra predicted from models composed of plane parallel layers of gas and those composed of spherical layers of gas were prompted by a desire to understand: (1) why the energy distributions of many O and Wolf-Rayet stars were "yellow," suggesting color temperatures of the order of 8000 to 12000 K, while the lines which were identified in the spectra of O and Wolf-Rayet stars suggested that one was observing an atmosphere in which the electron temperatures were greater than 25000 K; (2) why emission lines appeared in the spectra of Wolf-Rayet stars and of some O stars; and (3) why the absorption lines of O stars were rather broad with relatively small central absorptions, whereas the lines in stars of later spectral types were relatively deep. The early work using spherical model atmospheres is reviewed in Chapter 6 of *B Stars With and Without Emission Lines* (Underhill and Doazan, 1982), and this information will not be repeated here.

By 1940, it was apparent that most O and Wolf-Rayet stars were strongly reddened by interstellar extinction. After correction for this factor, the energy distributions of O stars in the visible range correspond to quite high temperatures. Nevertheless, even now, the intrinsic color indices for O stars are not well known, and some disturbing evidence remains, suggesting that some O stars either have intrinsic colors less blue than those of some early B stars, or that these O stars suffer a small amount of circumstellar reddening. See Part I for a review of this subject. A similar problem remains for the Wolf-Rayet stars. The intrinsic colors of Wolf-Rayet stars, so far as they can be determined, do not correspond to as high temperatures as are suggested by the lines which are seen in the spectra of Wolf-Rayet stars. Some Wolf-Rayet stars have unusually strong infrared excesses which extend into the visible range (see Underhill, 1980a). This may be why the measured color indices of some Wolf-Rayet stars are rather red.

In recent years, it has become evident from observations of the ultraviolet spectra of O stars that outflow occurs from most O stars. Ever since 1929, outflow has been recognized to be an important characteristic feature of Wolf-Rayet stars. The presence of high electron temperatures in the outer atmospheres of Wolf-Rayet stars has been obvious since the chief lines in the spectra of Wolf-Rayet stars were identified in the 1920's. Similarly, the ultraviolet spectra of O stars indicate that unusually high electron temperatures occur in some parts of the atmospheres of O stars. The detection of X rays from some O and Wolf-Rayet stars confirms unmistakably that electron temperatures of the order of 10^6 K exist in some parts of O-type atmospheres (Harnden et al. 1979; Sanders et al. 1982).

Because of this evidence for superheating and for outflow, as well as because of the evidence for inhomogeneity in the atmospheres of hot stars, we have taken the point of view that it is desirable to think of an O- or Wolf-Rayet-type atmosphere as being composed of two parts: a photosphere which is the place of origin of most of the continuous spectrum, and a mantle which radiates a spectrum which displays the results of the deposition of nonradiative energy and the effects of outflow. We have seen above that the photosphere of an O star can be modeled rather well by traditional methods.

In this section, we shall review the chief predictions of continuous spectra from model atmospheres having spherical geometry. The spherical model atmospheres which have been studied can be thought of as simple approximations to models for mantles. In most studies of spherical atmospheres, however, outflow is not taken into account. In few of the published studies is the deposition of nonradiative energy in the outer layers of a star considered.

We have emphasized above that it is one thing to make a model atmosphere and predict its spectrum following reasonable prescriptions for calculating the density and temperature structure and for solving the equations of radiative transfer, but that it is quite another matter to identify a particular model atmosphere with a real star. In the work which we shall report below, the authors have usually identified their model atmospheres with specific stars or spectral types by means of adopted values for the mass, photospheric radius, and effective temperature of the model star. Each of these quantities is poorly known for most O and Wolf-Rayet stars. The published identifications of the spherical model atmospheres with particular stars should be accepted with caution, because it is evident from the results of comparing the spectra from traditional model atmospheres with the spectra of real O and Wolf-Rayet stars that the intensities of the *lines* used to define O and Wolf-Rayet types are strongly influenced by conditions in the mantle of the star, and little by conditions in the photosphere of the star. This is important because only the conditions in the *photosphere* of the star are fixed by adopting a mass, radius, and effective temperature for the star. Conditions in the mantle are sensitively dependent on the amount of heating and outflow that occurs in the mantle. Since none of the static spherical model atmospheres made so far include the effects of nonradiative heating, none of them are truly acceptable as models for the mantles of O and Wolf-Rayet stars.

The model-atmosphere results reported in this section represent a triumph of ingenious numerical procedures. Some very difficult computing problems have been solved with great skill in order to obtain these results. Nevertheless, the physical situations which have been investigated are by no means truly representative of conditions in the mantles of O and Wolf-Rayet stars. Therefore, we advise against assuming that at present we truly understand what is going on in the atmospheres of O and Wolf-Rayet stars. One specific question to which we will return when all the results have been presented is whether spherical model atmospheres are really needed to "explain" the observed shape of the continuous spectra of O and Wolf-Rayet stars in the ultraviolet and visible ranges. There is little doubt that they are needed to "explain" the infrared excesses and radio emission which are observed for some O and Wolf-Rayet stars. We feel that it is important to keep open the point of view expressed above concerning using models composed of photospheres and mantles as an alternative to the option of attempting to explain the full spectrum of an O or Wolf-Rayet star by means of one model atmosphere having properties like those which have been investigated so far for plane parallel layer and spherical model atmospheres.

B. Physical Constraints

The geometrical configuration of the model atmospheres to be described in the next sections consists of spherical homogeneous shells of gas surrounding a spherical core for the star. The radiation field from this core comes from the center of the star, where it is generated by nuclear reactions. Its shape and intensity are described either by a Planck function at a prescribed temperature or by the continuous spectrum which would emerge from a traditional model atmosphere having prescribed values for the effective temperature and log g.

The composition of the models is described by specifying a value for the parameter $y = N(\text{He})/N(\text{H})$ and for selected values of y(light element) $= N$(light element)$/N$(H). Although the degree of ionization of the constituents of the atmosphere may change as a function of the radius, r, of the shell under consideration, the composition remains constant throughout the atmosphere. One of the necessary tasks when making a spherical model atmosphere is to find the distributions over possible ionization states (partition functions) for the various constitutents in each layer of the model.

The constituents of the model atmosphere are represented by means of model atoms, each of which is described by giving the energies and spectroscopic descriptions of its energy levels and the radiative and collisional probabilities for making transitions from one state to another in the model atom. The occupation numbers for the various energy states may be calculated from the local values for temperature and electron density by means of the Saha and Boltzmann laws. In this case, the condition of local thermodynamic equilibrium, LTE, is said to exist. On the other hand, one may insist that a state of statistical equilibrium exists for the populations of all the energy levels of the model atoms. Then non-LTE is said to exist. To find the non-LTE partition functions for model atoms, it is necessary explicitly to take account of the probabilities that atoms/ions may go from one energy state to another as a result of: (1) spontaneously emitting radiation, (2) a downward or an upward transition being caused by the presence of photons having an energy equal to the energy separation between the states, and (3) collisions with electrons. When non-LTE physics is used, a complex iterative method must be followed to determine the radiation field and the partition functions as functions of depth in the model atmosphere. The necessary procedures for solving this problem for a static spherical model atmosphere have been developed by Mihalas and Hummer (1974) from the complete-linearization method valid for planar atmospheres which was established by Auer and Mihalas (1969, 1972). In the case that the partition functions have been found by assuming LTE to exist, the equations for radiative transfer and for hydrostatic equilibrium may be solved by following the method described by Cassinelli (1971), that described by Castor (1974a), or that of Hundt et al. (1975). In the case of spherical model atmospheres in which outflow occurs, the equation of hydrostatic equilibrium is replaced by another equation (see *Characteristic Features of the Stellar Wind Problem*).

In the work with spherical model atmospheres, the coefficients for continuous absorption and for absorption in lines are often expressed in the form valid for single atoms or ions. Then these coefficients are multiplied by the relevant particle densities to obtain the opacity coefficient, $\varkappa_\nu$ or ℓ_ν, per unit volume in the model. The opacity due to electron scattering is evaluated by means of the Thomson scattering coefficient. Electron scattering is treated as though it is isotropic and coherent. During the part of the computational method of Mihalas and Hummer (1974) in which the non-LTE partition function for each constituent is found, the lines are assumed to have a Doppler profile. When the emergent spectrum from the model atmosphere is calculated, the line profiles are determined using the populations found for each level in the model, and full account is taken of the broadening of the line by Stark effect and by other causes.

Mihalas and Hummer (1974) have noted that, when one is faced with solving the equations of radiative transfer in spherical geometry (in order to apply the constraint of radiative equilibrium) and with simultaneously solving the equation of hydrostatic equilibrium (in order to ensure that the constraint of hydrostatic equilibrium is met), there is a fundamental conflict between the mathematically preferred depth scales for each set of equations. They chose to retain the radius, r, as the independent variable, using a Rosseland mean absorption coefficient to express distance in terms of an optical depth and to cope with the numerical instabilities which this action causes when the equation for hydrostatic equilibrium is solved. In the planar case, a similar computational difficulty may be avoided by replacing the radius with the column mass density as the independent variable. However, in the case of static spherical model atmospheres, a change to the variable $dm = \varrho\, dr$ is not advantageous.

When spherical model atmospheres are made, explicit account is taken of the variation of the attractive force due to gravity with the radius of each layer of the model. One uses $g(r) = GM_*/r^2$, and not $g = GM_*/R_*^2$. Here R_* is the nominal radius of the star and G is the gravitational constant. The second expression is the form which is used when making planar model atmospheres.

In order to obtain a significant extention of the model atmospheres, it is necessary to ensure that the effective value for gravity, g_{eff}, is small, but positive, at all values of r. The definition of $g_{\text{eff}}(r)$ is

$$g_{\text{eff}}(r) = g(r) - f(r)/\rho(r), \qquad (7\text{-}8)$$

where $f(r)$ is a force which acts outwards along the radial direction to diminish the attractive force due to the mass of the star. Experience has shown that, when $f(r)$ is made equal to the force exerted by the radiation which is absorbed in continuum wavelengths, i.e., when one puts

$$f(r) = \frac{\pi}{c} \rho \int_0^\infty \kappa_\nu F_\nu(r)\, d\nu, \qquad (7\text{-}9)$$

the effective extention of the atmosphere $\Delta r/R_*$ remains small. Here Δr is the distance between the layer where the radius is R_* and that where the optical depth is essentially zero. In Equation (7-9), $F_\nu(r)$ is the net flux of radiation at r in frequency ν; it is defined, as usual, in terms of the second moment of the specific intensity $I_\nu(r, \theta, \phi)$ at each point (r, θ, ϕ) in the spherical atmosphere. The quantity, $\varkappa_\nu$, is the monochromatic continuous absorption coefficient corrected for stimulated emission per gram of star material at r. The appropriate functional form for the force applied by radiation to each element of volume in the atmosphere is discussed further below.

Mihalas and Hummer (1974) found that they had to increase $f(r)$ arbitrarily from the value given by Equation (7-9) in order to obtain significantly extended spherical model atmospheres in radiative and hydrostatic equilibrium. They introduced an arbitrary radiation-force multiplier,

$$\gamma = \gamma_1 + \gamma_2 \exp(-\tau_R). \qquad (7\text{-}10)$$

Here γ_1 and γ_2 are assigned constants and τ_R is the Rosseland mean optical depth at r. Mihalas and Hummer (1974) and Kunasz et al. (1975) have explored the effects of different choices for γ_1 and γ_2. Truly extended model atmospheres are found only when γ takes artificially large values.

Spherical model atmospheres are defined by giving the composition in terms of fractional parts relative to the numerical abundance of hydrogen, the mass, M, and the luminosity, L, of the model star in absolute units and the radius, R, in linear units. That radius is at a specified optical depth, τ_R. An effective temperature for the model atmosphere may be found by evaluating the expression,

$$\frac{L_*}{L_\odot} = \left(\frac{R_*}{R_\odot}\right)^2 \left(\frac{T_{\mathrm{eff}}}{T_\odot}\right)^4, \qquad (7\text{-}11)$$

at the radius where $\tau_R = 2/3$. Alternatively, one may integrate the calculated emergent spectrum and use Equation (7-1) to find an effective temperature. A representative value for log g, useful for comparison with the values adopted for planar model atmospheres, may be determined by finding the value of $g(r)$ at the radius where $\tau_R = 2/3$.

The boundary conditions on the radiation field in spherical model atmospheres are: (1) that no radiation be incident at the surface layer where $r = r_{max}$, and (2) that at great depth (i.e., in a layer where τ_R is large), the radiation field has properties similar to those found for planar model atmospheres or that it can be described by a Planck function. In the case of spherical model atmospheres, a prescription must be written for determining the density at the effective outer radius of the model. Most investigators do this by postulating that the opacity due to electron scattering evaluated from infinity to the effective outer boundary of the atmosphere be equal to some selected small number. Because the equation for hydrostatic equilibrium is a first-order differential equation, this condition, together with specifying a value for the radius at $\tau_R = 2/3$ (or some other value), serves to define the solution of the equation for hydrostatic equilibrium. In the models of Mihalas and Hummer (1974) and of Kunasz et al. (1975), the radiation-force multiplier has a prescribed dependence on τ_R. This, with the above boundary conditions, defines the model atmosphere.

Alternatively, one may proceed as Cassinelli (1971) did. He chose to specify r_{max} and r_{min} in linear measure and to require that the total optical thickness of the atmosphere be a prescribed amount. The optical depth at r_{max} is set to a small value. These steps constrain the radiation-force multiplier to lie within a certain range. An appropriate value, independent of depth, is found by trial and error. Castor (1974a) studied what values of L and R were implied when g_{eff} had a range of small values, while Hundt et al. (1975) and Schmid-Burgk and Scholz (1975) attempted to find the regions in the HR diagram where static extended atmospheres might be expected to occur.

C. Properties of Hot Static Spherical Model Atmospheres

The methods for computing static spherical model atmospheres have been reviewed above, where it was noted that, because explicit account is taken of the variation of $g = GM_*/r^2$ with radius in the stellar atmosphere, a convention must be adopted concerning the definition of the equivalent values of T_{eff} and log g for comparison with models composed of plane parallel layers of gas in which g is assumed to be constant. Frequently the equivalent T_{eff} and log g are found by taking the values valid for the radius where the Rosseland mean optical depth in the atmosphere is 2/3. Another way of estimating the equivalent effective temperature of a spherical model atmosphere is to integrate its emergent spectrum.

D. Comparison of Planar and Spherical Model Atmospheres

We shall illustrate the differences in model structure and continuous spectrum resulting from the use of spherical geometry in the place of plane parallel layers by comparing nearly equivalent models selected from the sets of static non-LTE spherical model atmospheres of Kunasz et al. (1975) and of static planar models of Mihalas (1972b). We compare Model 16 of Kunasz et al. with the NLTE/L models of Mihalas having T_{eff} = 40000 and 50000 K, with log g = 4.0. Model 16 of Kunasz et al. has an artificially extended atmosphere created by using an arbitrary radiation-force multiplier (see Equation (7-10)) of $1 + 2.65 \exp(-\tau_R)$. Its defining parameters are $L/L_\odot = 6.17 \times 10^5$, $M/M_\odot = 60$, and $R/R_\odot = 10.94$. The boundary temperature is 35500 K, while the effective temperature found from the integrated flux is 38700 K. The equivalent effective temperature found from Equation (7-11) is 48900 K. From an observational point of view, the effective temperature of a star or a model atmosphere should be determined from the integrated flux. Consequently, the planar model with T_{eff} = 40000, log g = 4.0 provides the closest comparison between nearly equivalent planar and spherical model atmospheres.

The particle density-temperature structures of the selected model atmospheres are shown in Figure 7-15. Here temperature has been plotted against log N_{tot}. An X has been placed on each curve at about the position where the visible continuous spectrum is formed (i.e., at τ_R = 2/3 or mass = 1.0 g cm^{-2}). The continuous spectrum is formed deep in each model atmosphere in layers in which the total particle density, N_{tot}, varies between 10^{14} and 10^{15} cm^{-3}. The information reviewed in *Models of Static Plane Parallel Layers and Their Spectra* tells us that the characteristic lines by means of which O-type spectra are classified are formed in the layers of planar models in which the

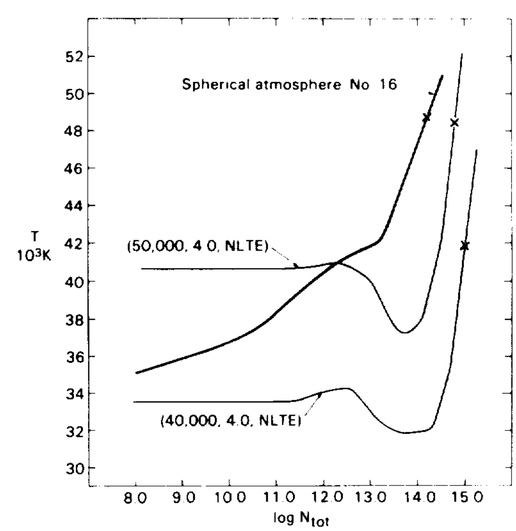


Figure 7-15. The run of temperature with particle density in equivalent non-LTE model atmospheres computed with planar and with spherical geometry. An X is placed at about the depth at which the continuous spectrum in the visible range is formed.

c-5

temperature is essentially constant and log N_{tot} varies between about 10 and 13.

The chief difference between a non-LTE, spherical model atmosphere and one composed of plane parallel layers is that the temperature continues to decrease outward through the line-forming region of a spherical atmosphere, while it remains nearly constant in a planar model. The gradient of temperature with particle density in the continuum-forming regions of all three model atmospheres is about the same.

The *shape* of the continuous spectrum in the observable range is compared for all three models in Table 7-3. Here Δm_{ν} from the intensity at $\lambda^{-1} = 1.80\ \mu m^{-1}$ ($\lambda = 0.5556\ \mu m$) is listed as a function of the reciprocal wavelength. These data have been taken from Kunasz et al. (1975) (KHM75) and Mihalas (1972b) (M72b). Between 1215 and 8201 Å, the spectra from Model 16 and the (40000, 4.0) model do not differ by more than ± 0.05 mag. For all practical purposes, the shapes are the same. The spectrum for model (50000, 4.0) is a little brighter at short wavelengths than that of either of the other models. This is to be expected because its effective temperature is significantly higher than that of either of the other model atmospheres. At wavelengths greater than about 1 μm, the spherical model atmosphere is relatively brighter than the two planar models when the energy distributions from all three models are normalized to the flux at 0.5556 μm. However, the spectrum of Model 16 of Kunasz et al. (1975) is not brighter than those of the planar models on an absolute scale, because the Eddington monochromatic flux, H_{ν}, at $\lambda^{-1} = 1.7559\ \mu m^{-1}$ is 2.241×10^{-4}, 4.86×10^{-4}, and 6.07×10^{-4} erg cm^{-2}s^{-1} Hz^{-1} respectively for the three models. These numbers will give the relative brightness of stars having the same effective photospheric radius at $\lambda^{-1} = 1.80\ \mu m^{-1}$ and lying at the same distance. The surface brightness of Model 16 in the continuum in the visible spectral range is only about 0.46 times that of a comparable planar model with $T_{eff} = 40000$ K.

Table 7-3
Shape of the Continuous Spectrum from One Spherical and Two Planar Model Atmospheres[a]

λ^{-1} (μm^{-1})	λ (μm)	Model 16 NLTE KHM75	(40000, 4.0) NLTE M72b	(50000, 4.0) NLTE M72b
8.231	0.1215	−2.008	−2.044	−2.178
6.554	0.1526	−1.817	−1.856	−1.972
4.877+	0.2050−	−1.494	−1.514	−1.610
4.877−	0.2050+	−1.474	−1.547	−1.622
3.810	0.2624	−1.139	−1.191	−1.254
2.744+	0.3645−	−0.617	−0.647	−0.690
2.744−	0.3645+	−0.680	−0.758	−0.768
2.401	0.4166	−0.473	−0.525	−0.532
2.058	0.4860	−0.223	−0.247	−0.250
1.756+	0.5695−	+0.043	+0.049	+0.049
1.756−	0.5695+	+0.048	+0.046	+0.047
1.524	0.6561	+0.294	+0.316	+0.321
1.219+	0.8201−	+0.698	+0.752	+0.764
1.219-	0.8201+	+0.684	+0.734	+0.750
0.953	1.0498	+1.134	+1.228	+1.249

[a]The magnitude difference between F_{ν} at λ^{-1} and F_{ν} at $\lambda^{-1} = 1.08\ \mu m^{-1}$ is listed.

Reference to the full monochromatic continuous spectra published by Kunasz et al. (1975) and by Mihalas (1972b) shows that the fraction of the total energy of the star which is radiated shortward of the Lyman limit in the case of a spherical model atmosphere is very much larger than it is for a planar model atmosphere. This happens because, in very opaque parts of the continuous spectrum, the spherical model atmosphere has a much larger effective radius than it has in relatively transparent parts of the spectrum, such as in the visible continuum. In the case of planar model atmospheres, the effective radius remains the same at all wavelengths.

The spherical model atmospheres of Kunasz et al. (1975) have been artificially extended by introducing an arbitrary radiation-force multiplier. Therefore, the differences between energy distributions from them and from equivalent planar model atmospheres should be a maximum. The energy distributions from spherical model atmospheres in which no arbitrary radiation-force multiplier has been used and those from planar model atmospheres having the same T_{eff} and log g have been compared by Gruschinske and Kudritzki (1979). These authors find that the differences in monochromatic magnitude when the two energy distributions are matched in the visible range are even smaller than the values shown in Table 7-3. One may conclude that the continuous spectra from spherical model atmospheres for O stars have essentially the same shape as the continuous spectra from planar atmospheres having the same effective temperature and effective log g values.

The extension of spherical model atmospheres as a function of wavelength has been investigated by Kunasz et al. (1975) and by Gruschinske and Kudritzki. This property of the spherical model atmospheres may be quantified by finding $r(t_\nu = 2/3)/r(\tau_R = 2/3)$, where τ_R is the Rosseland mean optical depth. Typically for Model 16, the extension is small. Except in very opaque parts of the continuous spectrum, it is less than 1 percent; only in the most opaque regions of the spectrum does it become as large as 10 percent. This means that the angular diameter subtended by a star which has been represented by such a model atmosphere will vary little with wavelength.

So far as interpreting the continuous spectra of O stars is concerned, the spectra and properties of planar model atmospheres appear to provide an adequate representation. Both Castor (1974a) and Schmid-Burgk and Scholz (1975) have considered the properties of LTE spherical model atmospheres for O stars. They conclude that Of stars probably have compact photospheres which can be represented reasonably well by means of planar model atmospheres.

It is only when one is interpreting the profiles of strong lines that spherical model atmospheres may be required. Then, however, systematic outflow is probably important. Consequently, static spherical model atmospheres are of little value for analyzing the spectra of O stars.

The profile of Hα in Model 16 of Kunasz et al. (1975) and that of Hα in the NLTE/L model (40000, 4.0) of Mihalas (1972b) are compared in Figure 7-16. In a spherical model atmosphere, the wings are made less deep, while the emission core is strengthened and the central part of the core absorption is deepened in comparison to what happens in an equivalent planar model atmosphere.

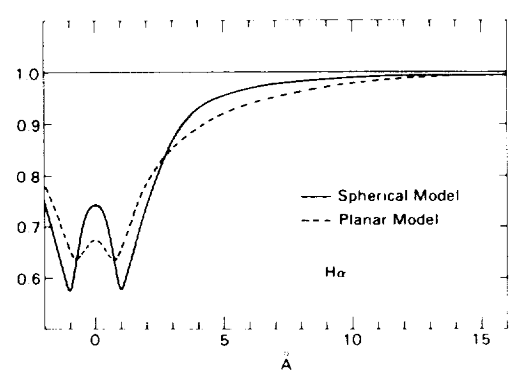


Figure 7-16. The Hα profile from a spherical non-LTE model atmosphere and from an equivalent planar non-LTE atmosphere.

E. Conservation Laws

The equation of state in a model atmosphere is that for a perfect gas. Then the following relationships exist between the partial pressure, P_i, of a constituent, i, its number density, N_i, and the isothermal speed of sound, a_i, in a gas composed of particles of type i:

$$P_i = N_i kT = \rho_i a_i^2 \qquad (7\text{-}12)$$

Here k is Boltzmann's constant and ϱ_i is the density of constituent i in the atmosphere. We recall that $\varrho_i = N_i m_i$ and $a_i = (kT/m_i)^{1/2}$, where m_i is the mass of a particle of type i.

We shall be discussing the special case of radial flow at a velocity, $v(r)$, in a stellar atmosphere. The velocity $v(r)$ is directed along a radius, and it is experienced by each element of volume in a spherical shell having a radius r and thickness dr. The velocity $v(r)$ is believed to arise as a result of the particles in the shell, (r, dr), acquiring an excess of momentum in the radial direction. We are interested in determining the source of this momentum and in determining how the resulting motion affects the spectrum of the star.

Because normal stellar atmospheres are composed mostly of hydrogen, it is convenient to use the isothermal speed of sound in hydrogen as a measure by which to characterize the flow motion. The flow is said to be subsonic when $v(r)$ is less than $a(r)$ for hydrogen at the ambient temperature and to be supersonic when $v(r)$ is greater than $a(r)$ for hydrogen. For orientation, the isothermal speed of sound in hydrogen at temperatures of interest for the atmospheres of O and Wolf-Rayet stars is given in Table 7-4.

Application of the principle that mass is conserved, none just "disappearing", results in the equation of continuity:

$$\frac{\partial \rho}{\partial t} + \nabla \cdot \rho \mathbf{u} = 0. \qquad (7\text{-}13)$$

Here we are representing the material in the atmosphere as a fluid with density ϱ and a three-dimensional velocity, $\mathbf{u}$. The first term in Equation (7-13) represents fluctuations in the density that depend only on the time.

Application of the principle of the conservation of momentum and use of the equation of state results in the generalized equation of motion:

$$\frac{\partial \mathbf{u}}{\partial t} + \mathbf{u} \cdot \nabla \mathbf{u} = -\frac{1}{\rho} \nabla(\rho a^2) + \mathbf{F} + \mathbf{f}. \qquad (7\text{-}14)$$

Here $\mathbf{F}$ is the net body force acting on the material in the stellar atmosphere due to the star itself. It may consist of an inward-directed gravitational force due to the mass of the star and an outward-directed force due to radiation, as well as forces due to momentum transfer from waves, or forces associated with velocity fields originating beneath the photosphere. If magnetic fields are present and the material in the atmosphere is in motion, magnetohydrodynamic (MHD) forces may act. The term $\mathbf{f}$ represents forces exerted on each element of the atmosphere of the star by the environment of the star. Since at this time we are interested only in isolated single stars, we shall set the term $\mathbf{f}$ equal to zero. In the case of a binary star, this term would represent the tidal force due to the companion star. The term $\mathbf{f}$ may also be used to represent the pressure exerted by the interstellar medium on a star.

The principle of conservation of energy used with the equation of state leads to an equation describing the energy exchange which occurs in each element of volume in the model atmosphere. We have

$$\frac{\partial E_g}{\partial t} + \mathbf{u} \cdot E_g + \rho a^2 \frac{\partial}{\partial t}\left(\frac{1}{\rho}\right) + \rho a^2 \, \mathbf{u} \cdot \nabla \left(\frac{1}{\rho}\right) = G - L. \qquad (7\text{-}15)$$

Here E_g is the internal energy of the material in the atmosphere per unit volume, G represents the rate of gain of energy per unit volume, and L represents the rate of energy loss per unit volume. Each element of volume gains energy

Table 7-4
Velocity of Sound in Hydrogen as a Function of Temperature

T (K)	a (km s^{-1})	T (K)	a (km s^{-1})
1000	2.9	10000	9.1
2000	4.1	20000	12.8
3000	5.0	30000	15.7
5000	6.4	50000	20.3

from the radiation which enters it from the surrounding atmosphere, and it loses energy by sending radiation to the surrounding part of the atmosphere. An element of volume may gain energy as a result of magnetohydrodynamic events releasing in it energy which had been stored in motions in the photosphere (see Chapter 8). Also, an element of volume may gain energy as a result of the dissipation of shocks or of mechanical waves generated elsewhere in the star. The annihilation of magnetic flux may deposit energy locally. An element of volume may lose energy by generating shocks and mechanical or MHD waves. Under some circumstances, the conduction of energy into and out of the element of volume by electrons may produce significant contributions to $G - L$.

It is Equation (7-15) which enables us to find the temperature law, $T(r)$, in the atmosphere. In the case of traditional model atmospheres, radiative equilibrium is assumed to exist. In this case, energy is transferred only by means of radiation, and enforcing the condition that the luminosity of the star,

$$L = 4\pi^2 r^2 \int_0^\infty F_\nu \, d\nu, \tag{7-16}$$

be independent of the radius in the atmosphere enables one to find $T(r)$. In planar model atmospheres, the radius of each layer has the same constant value. This permits one to define the effective temperature of a star by means of Equation (7-1). In the case of spherical model atmospheres, the equivalent effective temperature is found from the luminosity of the star and an assigned effective radius for the photosphere or from the integrated emergent spectrum.

If a steady state is postulated to exist, and if only radial motion is considered, Equations (7-12), (7-13), and (7-14) result in the following equation for the radially directed component of velocity as a function of radius:

$$\upsilon(r) \frac{d\upsilon}{dr}\left(\frac{a^2}{\upsilon^2} - 1\right) = \mathbf{F}(r) - \frac{2a^2}{r}\left(1 - \frac{d \ln T}{d \ln r^2}\right). \tag{7-17}$$

Here $\mathbf{F}(r)$ is the net force acting in the radial direction on each element of volume in the shell (r, dr). In what follows, we shall review studies which are concerned only with radial forces and with radial flow. In particular, we shall not review what might occur if azimuthal and latitudinal forces were active due to rotation of the star and to the action of viscosity. Likewise, we shall not attempt to investigate the effects of MHD forces which act in a general direction. We do, however, note that MHD forces may have a component in the radial direction. It is clear from observations of the expulsion of spicules and jets in the case of the Sun that impulsive forces act in stars. One needs to postulate the presence of magnetic fields if MHD forces are to be active.

Equation (7-17) has the property that a singularity occurs when $\upsilon(r) = a(r)$. An appropriate solution for $\upsilon(r)$ in a stellar atmosphere is postulated to be one such that $\upsilon(r)$ and all its derivatives with respect to r are continuous at the singular point, and that, at great distances from the surface of the star, the moving material melds with the interstellar medium at a finite velocity. The radial component of velocity, $\upsilon(r)$, is usually assumed to be small, but directed outward, at the interface between the underlying stationary photosphere and the flowing atmosphere.

Use of Equation (7-17) with a temperature law $T(r)$ permits one to find the density and velocity structure in a flowing model atmosphere. When υ^2 is small with respect to a^2, say less than 0.1 a^2, one can usually ignore the effect of the motion on the density structure of the model atmosphere and use the equation of hydrostatic equilibrium to find the dependence of density and pressure on radius.

The mathematical properties of Equation (7-17) have been discussed extensively in the literature on solar and stellar winds. A review in the context of the solar wind is given in Chapter 15 of Mihalas (1978).

1. Special Cases Relevant to the Atmospheres of Hot Stars. No model-atmosphere studies of early-type stars have been made which retain the terms which describe the time variation of the variables in Equations (7-13), (7-14), and (7-15). In every study, it is assumed that the flow is steady. The partial derivatives with respect to time are put equal to zero.

For a one-dimensional steady flow, Equation (7-13) reduces to

$$r^2\ \partial(r^2\ \rho v)/\partial r\ =\ 0, \qquad (7\text{-}18)$$

which implies

$$4\pi r^2 \rho(r)v(r)\ =\ \text{constant}\ =\ \dot{M}. \qquad (7\text{-}19)$$

Here $\dot{M}$ is the rate of loss of mass from the atmosphere. This equation is used with the velocity and density laws deduced by means of flowing model atmospheres to estimate the rate of mass loss from stars.

Since Equation (7-19) must be valid at the sonic point in a wind, it is possible to estimate the radius of the sonic point in the case of an assigned value for $\dot{M}$. In the mantles of O-type supergiants, the temperature at the sonic point may be of the order of 30000 K, while the density may be of the order of 2×10^{-12} g cm^{-3}. These numbers are typical of the densities and temperatures in the outer parts of the model atmospheres of Kurucz (1979) and of Mihalas (1972b) for effective temperatures in the range 30000 to 35000 K and log g = 3.5. Recalling that the sonic velocity is

$$v_c\ =\ a\ =\ (kT/m_i)^{1/2}, \qquad (7\text{-}20)$$

where m_i is the mass of a typical ion in the flowing plasma, we find that the product $\varrho(r)v(r)$ at the sonic point may typically be expected to have a value near 3.2×10^{-6} g cm^{-2} s^{-1}. The rate of mass loss from an O-type supergiant is believed to be of the order of 10^{-6} $M_\odot$ yr^{-1}. (See *Rate of Mass Loss from Hot Stars* later in this Chapter for some estimated values of $\dot{M}$.)

These figures result in the following relation (Thomas, 1973) for the radius of the sonic point, r_c, of an O-type supergiant:

$$r_c/R_*\ =\ 56.9(v_c\rho_c)^{-1/2}\dot{M}^{1/2}R_*^{-1}\ . \qquad (7\text{-}21)$$

Here the product $v_c\varrho_c$ is to be entered in units of 3.2×10^{-6} g cm^{-2}s^{-1}, $\dot{M}$ in units of 10^{-6} $M_\odot$ yr^{-1}, and R_* in solar radii. For an O9 supergiant having a radius of 35 $R_\odot$, and with $v_c\varrho_c$ and $\dot{M}$ equal to unity in the selected units, the sonic point occurs at 1.6 stellar radii.

In the case of O-type main-sequence stars, $\dot{M}$ may be smaller by a factor of the order of 10^{-3} than it is for O-type supergiants, while the radius will be smaller by a factor of the order of 3. Since $(v_c\varrho_c)^{-1/2}$ is unlikely to change by a large factor when going to the main sequence, one sees that in main-sequence O stars, the sonic point may be expected to lie at a distance of about 1.15 stellar radii from the center of the star. The sonic point may be expected to move close to the photosphere for large stars such as the middle and late B-type supergiants because, although $\dot{M}$ becomes a little smaller for such stars than it is for O-type supergiants, $v_c\varrho_c$ will also become smaller. For O- and B-type supergiants, r_c/R_* may be expected to scale about as R_*^{-1}.

Equation (7-17) is used to find the density and velocity structure of flowing model atmospheres for hot stars. The cases which have been studied are reported below. In all of these applications, the net body force, **F**, acting on the material in the atmosphere is taken to be the vector sum of the force due to gravity and that due to radiation. The forms which are used for the force due to radiation are described in the next section. In this work, it is sometimes assumed that the temperature is constant, a not unreasonable assumption considering the results for static model atmospheres shown in Figures 7-3 and 7-4.

If the flow velocity in the model atmosphere is identically set to zero, the atmosphere is said to be in hydrostatic equilibrium. For a spherical model atmosphere, Equation (7-14) then reduces to

$$\frac{dP}{dr} = -\frac{GM\rho}{r^2} + f_{\text{rad}}\,\rho. \qquad (7\text{-}22)$$

Here we have used Equation (7-12) and we have introduced the functional form for the force of gravity. The dependence on radius of the force due to radiation is discussed in the next section.

In the case of a steady one-dimensional flow, Equation (7-15) becomes

$$v(r)\left[E_g + P\frac{\partial}{\partial r}\left(\frac{1}{\rho}\right)\right] = G - L. \qquad (7\text{-}23)$$

If there is no flow and if energy is transmitted only by radiation, Equation (7-15) reduces to the statement that radiative equilibrium should exist. If there is flow and if energy is brought into each element of volume and removed from it by means other than radiation, such as conduction, the dissipation of mechanical or MHD waves, or the occurrence of shocks, the appropriate forms must be used for G and L. An equation for determining the temperature law will then be obtained.

F. Mechanical Force Exerted by Radiation

This force is the *net* result of the atoms and ions in a stellar atmosphere *absorbing* photons from a stream of radiation which is flowing in a specified direction and emitting radiation in the same direction as a result of stimulated emission. Since spontaneous emission of radiation takes place isotropically, spontaneous emission does not affect the distribution of momentum among the particles in a stellar atmosphere. We recall that the net rate of transfer of momentum across a surface can be expressed as a force acting on the surface. Each photon of frequency ν which is absorbed by an atom or ion transfers a quantum of momentum equal to $h\nu/c$ to the atom or ion. The acquired quantum of momentum is directed in the direction of flow of the absorbed radiation. Likewise, each emitted photon takes from the atom or ion a quantum of momentum equal in amount to $h\nu/c$ and directed in the direction in which the photon is emitted.

The force exerted on each element of volume in a stellar atmosphere by the radiation field which is present can be found. Let the absorption coefficient at frequency ν due to one type of atom or ion be k_ν per atom in the atmosphere. The coefficient k_ν may represent absorption in continua or in lines. We note that the absorption of a photon lifts the atom or ion from state 1 to state 2, which has higher energy than state 1. For this transition, the Einstein probability for absorption is B_{12}, while that for stimulated emission is B_{21}. We postulate that the populations of levels 1 and 2 are in the ratio N_1/N_2. It may be readily shown (see, for instance, Chapter 5 of Chandrasekhar, 1939) that, if the monochromatic flux of radiation in the stellar atmosphere in the direction r is $F_\nu(r)$, then the force on the atoms/ions due to the absorption of radiation is

$$d f^i_{\text{rad}}(r,\nu)\,d\nu = \frac{\pi}{c}k_\nu \,(1 - N_2B_{21}/N_1B_{12})m_iN_iF_\nu(r)\,d\nu. \qquad (7\text{-}24)$$

Here m_i is the mass of type i of atom or ion, and the species has a number density of N_i in the layer of radius r. The term $(1 - N_2B_{21}/N_1B_{12})$ gives the correction for stimulated emission. When one integrates over the full frequency range, one finds the total force exerted on the atoms/ions of type i in each element of volume in the shell of radius r. It is

$$f^i_{\text{rad}} = \frac{\pi}{c}\int_0^\infty (k'_\nu + \ell'_\nu)_i N_i m_i F_\nu(r)\,d\nu. \qquad (7\text{-}25)$$

Here k_ν' has been written for the total continuous absorption coefficient corrected for stimulated emission due to all of the continua of atoms/ions of type i, and ℓ_ν' has been written for the summed line-absorption coefficient at frequency ν corrected for stimulated emission. One usually thinks in terms of only one line at any frequency ν, but of several possible spectra of continuous absorption. In Equation (7-25), F_ν stands for the net monochromatic flux at depth r in the atmosphere.

Electrons, protons, and alpha particles make up a large fraction of the particles in each element of volume in the atmosphere of a hot star. The force exerted by radiation on them cannot logically be evaluated by means of the arguments leading to Equation (7-25) because electrons, protons, and alpha particles do not actually absorb radiation. However, they are able to acquire momentum from the radiation field. The existence of the Compton effect demonstrates this. The cross sections for the interactions between radiation and these types of particle can be found by means of the Klein-Nishina formula. In the nonrelativistic regime, the case where $h\nu/mc^2 << 1$, which is the case valid for most stellar atmospheres, the interaction cross section reduces to the Thomson scattering cross section. Consequently, it follows that the maximum force experienced by an element of volume containing N_e electrons per unit volume is

$$f_{\mathrm{rad}}\ (\text{electrons}) = \frac{\pi}{c}\,\sigma_T N_e m_e \int_0^\infty F_\nu(r)\,d\nu, \tag{7-26}$$

where σ_T is the Thomson scattering coefficient for electrons. We have

$$\sigma_T = \frac{8\pi}{3}\left(\frac{e^2}{mc^2}\right)^2 . \tag{7-27}$$

Similar formulas may be written for the force acting on protons and alpha particles. However, the numerical values of these forces are negligible in comparison to the force on the electrons because the masses of protons and of alpha particles are so much greater than the mass of the electron.

According to Equations (7-25) and (7-26), each species in a stellar atmosphere experiences a different radially directed force as a result of the action of radiation. So far as its flow properties are concerned (see Equation (7-17)), the stellar atmosphere may be supposed to act like a collection of different fluids, because each constituent experiences a different radial force which is the vector sum of the force due to radiation and the force due to gravity. Here we are neglecting the action of radial forces which may originate because of electrostatic and MHD forces. The action of such forces should be considered explicitly. However, it has been neglected in the existing theories of the winds of hot stars. The extreme situation of having to treat the stellar atmosphere as a set of independent fluids will disappear if the individual particles exchange momentum as a result of collisions and coulomb interactions at a rate which is rapid compared to the rate at which they change their flow velocity. In all treatments of the winds of hot stars, it has been assumed, indeed, that a sufficiently rapid exchange of momentum takes place between the various constituents of the wind that one can write

$$f_{\mathrm{rad}}(r) = \frac{\pi}{c}\,\rho(r)\int_0^\infty (k_\nu' + \ell_\nu' + \sigma)F_\nu(r)\,d\nu. \tag{7-28}$$

Here k_ν', ℓ_ν', and σ are evaluated at each radius, r, per gram of star material using the usual formulas. It follows that the acceleration due to the action of radiation may be found from

$$g_{\mathrm{rad}}(r) = f_{\mathrm{rad}}(r)/\rho(r). \tag{7-29}$$

One can define an effective gravity which acts on each element of volume in the stellar atmosphere at each radius such that $g_{\mathrm{eff}} = g - g_{\mathrm{rad}}$.

The frequency of collisions needed in the atmosphere of an O star in order for the momentum acquired by individual species to be distributed among all species has been investigated by Castor et al. (1976). They show that, unless the density in a wind is very low, it is probably appropriate to use Equation (7-28) to define a value of g_{rad} which is applicable to all species. However, the subject deserves more study because exactly the opposite point of view is adopted by Michaud (1977) and his collegues in their theory of the origin of the abundance anomalies in Ap stars.

In the diffusion theory of the abundance anomalies in Ap stars (see, for instance, the review papers by Michaud, 1977, and Vauclair and Vauclair, 1977), specific account is taken of the force exerted by radiation on *particular* species in order to obtain an anomalous composition. It is concluded that if turbulence is small enough, a *differential* radiative force will act with electrostatic forces to provide a separation of elements in times of the order of 10^4 years. More information about the Ap stars can be found in Wolff (1983).

The temperatures which are considered in the layers of model Ap stars where diffusion is predicted to occur are similar to those in the outer atmospheres of the O and B stars which are observed to show outflow. However, the densities in the layers of the model Ap stars where diffusion is believed to occur are probably *higher* than those in the flowing parts of the atmospheres of O and B stars. Consequently, it is an unanswered question why in the one case—winds from hot stars—one is justified in assuming that momentum acquired from the radiation field by one species is redistributed to all species present by collisions, while in the other—the subphotospheric layers of Ap stars—one does not have to make this assumption.

Diffusion is not expected to be effective in O and B stars with winds because the atmospheres of these stars are believed to be turbulent and because most of the O and B stars are known to rotate fairly rapidly. Rapid rotation implies substantial motion in the subphotospheric layers. The question appears to be open whether or not the driving force exerted by radiation on individual atoms or ions is, indeed, effectively distributed to all species in the atmosphere. If this is true, then the diffusion theory for abundance anomalies in Ap stars should be examined again. If it is not true, then the possibility must be considered that the winds from O and B stars do not have cosmic composition.

Equation (7-28) has been evaluated at each level of the atmosphere by Kurucz et al. (1974) for their LTE line-blanketed model atmospheres. Some of these model atmospheres have effective temperatures and log g values in the range which is believed to be of interest for O stars. In every case, g_{rad} is less than g because the values of log g have been chosen so that this will occur. On the whole, in the parts of the hottest models where the strong lines are formed, the ratio,

$$\Gamma = g_{rad}/g,$$

falls in the range 0.1 to 0.5. In the deep layers of some of the models, Γ approaches 0.8 to 0.85. Similarly, in the line-blanketed model atmospheres of Kurucz (1979), Γ is rarely larger than 0.5 in the layers in which the emergent line spectrum is determined (see Kurucz and Schild, 1976). These numerical results indicate that, in normal stellar atmospheres for which the particle densities and electron temperatures may be expected to be about as in the model atmospheres of Kurucz and his colleagues, Γ is usually less than unity by a significant factor.

The question of how to represent consistently the contribution of a strong line to g_{rad} has been examined by Castor (1974b). He explored three cases: (1) that for a two-level atom without a continuum, (2) that for a two-level atom with a continuum present, and (3) that in which line absorption takes place in an atmosphere in which there are large velocity gradients. The line-formation process is treated as noncoherent scattering with a probability ϵ that an atom which has been excited to its upper energy state by the absorption of a photon will not decay radiatively. For resonance lines in stellar atmospheres, ϵ may lie in the range 10^{-6} to 10^{-8}.

The result in case 1 is that the force due to the radiation varies as follows:

$$f_{rad}(\text{line}) \approx \frac{\kappa_L \Delta\nu_D}{c} \epsilon^{1/2} F(\tau, \phi(\nu)). \quad (7\text{-}30)$$

Here $\varkappa_L$ is the value for the absorption coefficient at the center of the line and $\Delta\nu_D$ is the thermal Doppler halfwidth for the line. In Equation (7-30), $F(\tau, \phi(\nu))$ is a function of the optical depth in the line center, τ, and its form depends upon the shape function for the line, $\phi(\nu)$. Castor (1974b) gives expressions for $F(\tau, \phi(\nu))$ for the cases of Doppler and of Voigt profiles. Because ϵ is small for resonance lines, the term $\epsilon^{1/2}$ is the dominating factor in determining the value for f_{rad}(line). Its effect is to make f_{rad}(line) very small in the case of resonance lines with a negligible continuum opacity. Castor also notes that the source function in a resonance line is small, varying as $\epsilon^{1/2}$. Consequently, there is very little radiation present at frequency ν to interact with the atoms and create a force due to radiation.

Castor obtained similar results for case 2, which treats the situation in which a continuum is present. The presence of continuous absorption increases the probability that photons are destroyed (absorbed) rather than just being redirected (scattered). The result is that f_{rad}(line) is larger than for case 1. It is still rather small for resonance lines since, in the case of a Doppler profile, it is proportional to the ratio, β, of the continuous absorption coefficient at ν to the absorption coefficient in the center of the line. In the case of a Voigt profile, the proportionality factor is $(a\beta)^{1/2}$, where a is the damping parameter for the Voigt function.

Castor concludes by stating that the force of radiation due to absorption in a strong line in a stationary atmosphere can be written as follows:

$$f_{rad}(\text{line}) \approx \frac{4\pi}{c} \kappa_L \Delta\nu_D B \min(1, 1/\tau). \quad (7\text{-}31)$$

Here τ is the optical depth in the center of the line, and it is generally large. The quantity B is the Planck function evaluated at the frequency of the line center for a temperature which is characteristic for the line-formation layers.

In the third case studied by Castor, a steep velocity gradient is supposed to be present along the line of sight in the line-forming region of the atmosphere. This effectively decouples each element of the atmosphere from neighboring elements and divides the resonance line into a continuous set of optically thin absorption lines, each displaced by the Doppler effect with respect to the zero line-of-sight velocity position for the line. The line-absorbing atoms in the moving atmosphere are exposed to a diffuse field of radiation composed of line photons which have been scattered and to an attenuated direct field of radiation from the core of the star. Castor shows that the latter field of radiation is the important one for transferring momentum to the atoms. He concludes that in case 3 one can write:

$$f_{rad}(\text{line}) \approx \frac{1}{c} \kappa_L \Delta\nu_D F_c \min(1, 1/\tau), \quad (7\text{-}32)$$

where F_c is the local continuum flux and τ is a modified optical-depth variable such that

$$\tau = \kappa_L \rho v_{th}/(dv/dz)_{AV}. \quad (7\text{-}33)$$

Here $(dv/dz)_{AV}$ is a typical value for the velocity gradient in the atmosphere along the line of sight. The optical depth τ is a local quantity which corresponds to the number of atoms in a cylinder whose height is the distance in which the velocity changes by one thermal unit. Because the value of τ which is determined by Equation (7-33) is much smaller than the value of τ which enters Equation (7-31), the force due to the absorption of radiation by a resonance line in an expanding atmosphere may become large.

The question of supplying momentum from sources other than radiation is discussed in Chapter 8, *Momentum Transfer*.

G. Characteristic Features of the Stellar Wind Problem

When we consider the theory of stellar winds, we must work with a different conception of the forces acting in a stellar atmosphere, of the sources from which the observed energy in the stellar atmosphere may come, and of the boundary conditions imposed on the model star than the one which has provided a useful basis for the traditional theory of stellar atmospheres and the traditional theory of stellar structure and evolution. The change in thinking which is required has been described by Pecker et al. (1973) and by Thomas (1973). These investigators emphasize that the stellar atmosphere is a transition zone between the stellar interior and the interstellar medium. Any acceptable solution for a model atmosphere must be such that the gradients of density, temperature, and motion, as well as the absolute values of these quantities, match smoothly with the values in the interior of the star and with the values in the surrounding interstellar medium. Pecker et al. note that one should think in terms of storage modes for mass and for radiation, as well as in terms of fluxes of mass and radiation.

The equations which will describe the situation are those which can be derived from the conservation equations. When attempting to understand the significance of stellar winds, one must consider carefully which terms to retain in the most general formulations of the conservation equations and which it is permissible to neglect. One also must consider what the appropriate boundary conditions are for the problem in hand.

The need for taking such steps is not new. The traditional equations and boundary conditions which underlie the traditional theory of stellar atmospheres reviewed at the beginning of this chapter were adopted as reasonable for understanding a major fact about stars—the fact that the intrinsic properties of stars change very slowly relative to the time scales for most physical processes known on the Earth. In order to make models of pulsating stars, it was necessary to think through anew the question of what terms to retain in the conservation equations and what boundary conditions to set so that meaningful solutions might be obtained. The steps in this procedure are clearly outlined in the review article by Ledoux (1965) and in that by Ledoux and Walraven (1958). When pulsation and its change is a significant observable fact to be explained, one must think carefully about the kind of stability which occurs in the part of the model star which is under consideration. The same questions have to be considered once again as we begin to make theories for the winds from early-type stars.

In what follows, we shall outline the major statements that have been published in recent years concerning the factors which may constrain the winds from early-type stars. This involves deciding on the forces which act in a wind and the sources of energy which may be tapped in the wind. A succinct review has been given by Hearn (1981).

1. Three Theoretical Starting Points. Three approaches have been taken towards developing theories for the winds from early-type stars. The first has been to concentrate attention on the outward-directed force due to radiation, particularly that absorbed in lines, and either to assume that an isothermal situation exists or to allow the temperature law to be established by ensuring that radiative equilibrium occurs. The second approach has been to direct attention toward maintaining an elevated temperature in a region of the mantle just outside the photosphere, and to determining what can be deduced about the flow of a wind and a possible supply of energy from knowledge of the processes which are postulated to provide the required high temperatures. The third approach has been to concentrate attention on a possible source in the subphotospheric layers for the energy and momentum which is seen in the wind, and then to deduce what the consequences of the presence of this source may be for the characteristic properties of a wind. Additional factors are discussed in Chapter 8.

The first approach was introduced by Lucy and Solomon (1970), who developed ideas originating with Milne (1926) and expanded by Gerasimovic (1934). Lucy and Solomon directed attention to the force of radiation resulting from the absorption of radiation by resonance lines. The implications of the action of the force due to radiation have been discussed by Castor and Cassinelli (1973). Their deductions are summarized below. Model flowing atmospheres for hot stars resulting from the action of the force of radiation have been published by Cassinelli and Hartmann (1975) and by Castor et al. (1975). Although these authors identify their model atmospheres with specific types of stars, their identifications must be accepted with reserve because of the factors which are discussed in the section *The Physics of Winds*.

Cassinelli and Hartmann make models of the region of the atmosphere between the photosphere, where outflow is negligible, and the point at which the flow velocity equals the isothermal sound velocity. They assume that energy is put into the model atmosphere only by means of the radiation field and that radiative equilibrium exists. To obtain a model atmosphere of the desired extent, they are forced to introduce an arbitrary radiation-force multiplier. Castor et al. (1975) develop an analytical expression to represent the force due to the radiation absorbed in lines, and they demonstrate the consequences of the action of this force. Their results have greatest validity in the trans-sonic region of the wind. Castor et al. assume that radiative equilibrium exists and that no energy is put into the model atmosphere except by means of the radiation field. The theory by Castor et al. has been broadened by Klein and Castor (1978) to give an explanation of the emission seen in the spectra of O stars at Hα and some of the He II lines. More recently, Weber (1981) has reviewed critically the theories of winds driven by the force due to radiation which is absorbed in lines; he has produced some additional flowing model atmospheres. Abbott (1980, 1982) has studied further the theory of winds driven by the force due to radiation absorbed in lines and has elaborated the consequences of the adopted equation of motion for a wind.

The second approach has been followed by Hearn (1975a, 1975b, 1983). In his work, Hearn has explored the possibility that the mass-loss mechanism for early-type stars is the pressure imbalance which may result from the presence of a hot corona. Hearn has evaluated the effects of particular processes which may be at work to provide the hot corona which drives his winds. Additional detail has been worked out by Hearn and Vardavas (1981a, 1981b) and by Vardavas and Hearn (1981).

The third approach has been taken by Thomas (1973). Thomas infers that the presence of a wind implies the existence of a field of kinetic energy below the photosphere, and he deduces the constraints which are required in order to attain a stellar wind which has a pattern of flow and density like what is observed for early-type stars. However, although Thomas notes that pulsation and convection may occur in layers below the photosphere, he does not postulate a specific process for generating the needed field of kinetic energy. Rather he argues that the observation of mass loss from stars implies that such a field exists. His statements seem to suggest that at considerable depths in the star this field of kinetic energy can be described in terms of stochastic processes. Thomas works out the consequences of the presence of his inferred source of energy and momentum. He takes into account the action of the force due to radiation in the outer part of a stellar atmosphere, and he places the special cases studied by Lucy and Solomon (1970) and by Castor and Cassinelli (1973), (cases in which there is no energy source explicitly mentioned other than the radiation field), in the context of a general theory of stellar atmospheres in which directed motion occurs. Further details concerning a stellar wind powered by a stochastic source of kinetic energy are worked out by Cannon and Thomas (1977). Cannon and Thomas pay particular attention to the stability of the stellar atmosphere against small changes of the kinetic energy of the particles in the atmosphere. Andriesse (1978, 1980a, 1980b, 1981) has used the concept of a field of thermodynamic fluctuations which can be described in stochastic terms to derive an expression for the rate of mass loss from a star in terms of the mass, radius, and luminosity of the star.

The results of Lucy and Solomon, of Castor and Cassinelli, of Castor et al., of Cassinelli and Hartmann, of Weber, and of Abbott are obtained by postulating the action of one specific force, that exerted by radiation. These authors introduce no explicit sources of additional heating. The results of Hearn and of Hearn and Vardavas detail the implications of specific heating processes in the part of the atmosphere from which the wind originates. They downplay the action of the force of radiation. The results of Thomas, of Cannon and Thomas, and of Andriesse are not easy to accept because they do not identify the explicit sources of the departures from strict hydrostatic and radiative equilibrium in a stellar atmosphere. Thomas, Cannon and Thomas, and Andriesse have felt, with some justification, that at the time their studies were made, a particular force could not be surely specified for causing the departures from strict equilibrium conditions that the existence of radiation and mass flows imply. To work out the consequences of a stochastic distribution of the departures from equilibrium conditions is a valid procedure in theoretical physics. Because explicit forces and explicit sources of heating in the stellar wind are postulated in the theories of types 1 and 2, these theories have been more easily accepted than those of type 3.

A simple introduction to the basic principles which are involved in solving Equations (7-13), (7-14), and (7-15) simultaneously to obtain flowing model atmospheres can be found in Böhm (1973). The key statements of the theories for stellar winds, except the theory for mass loss by Andriesse, have been briefly summarized by Hearn (1979). The reasons why the observation of stellar winds forces us to change our viewpoint substantially from that which has sufficed for making traditional model atmospheres have been summarized by Thomas (1977, 1979). A review of the theories of the winds from early- and late-type stars has been presented by Cassinelli (1979).

2. Properties of Spherical Flow. The discovery of mass loss from O9.5 and B0 supergiants by means of moderate resolution spectra taken from a sounding rocket (Morton, 1967a, 1967b) and its confirmation by means of spectra of better resolution (Morton et al., 1968) directed attention toward the problem of spherical flow and how such a flow might be modeled using Equation (7-17). Studies of the solar wind have demonstrated the general properties of an acceptable solution to Equation (7-17). The radius in the atmosphere at which the flow velocity, $v(r)$, equals the ambient sound velocity, $a(r)$, divides the flowing atmosphere into two parts. The first is a subsonic region lying between the radius of some deep layer below which flow is negligible and below which the atmosphere may be represented by a model which is in hydrostatic equilibrium, and the radius at which the critical condition that $v(r) \equiv a(r)$ occurs. The second is a region in which the flow velocity takes on supersonic values, $v(r) > a(r)$. Acceleration to velocities in the range 2000 to 3500 km s^{-1} occurs in this region for O stars. In the case of O stars, the critical velocity is believed to lie in the range of 15 to 20 km s^{-1}.

An acceptable solution for $v(r)$ is one such that the function $v(r)$ is regular through the critical radius and such that $v(r)$ remains finite as r becomes large. The condition that $v(r)$ remain regular at the sonic radius, r_c, means that the right-hand side of Equation (7-17) must vanish when $v(r_c) = a(r_c)$. This puts a constraint on the value required for $\mathbf{F}(r)$ and on the temperature gradient which exists at r_c. Both $\mathbf{F}(r)$ and the temperature gradient should be continuous functions of radius.

If one supposes that the only forces acting on the gas are those due to gravity and to an outward force which is directed radially, one finds that an acceptable solution will be obtained at the critical point when

$$\frac{GM\rho_c}{r_c^{\,2}}(1-\Gamma_c) = \frac{2a^2}{r_c}\left(1-\frac{1}{2}\frac{r_c}{T_c}\left(\frac{dT}{dr}\right)_c\right). \tag{7-34}$$

Here a quantity Γ, which represents the ratio of the outward force to the inward force due to gravity, has been introduced. A subscript c is attached to the quantities which depend on radius to indicate that one is concerned with their values at the critical radius where the flow velocity has a value equal to the ambient sound velocity in the atmosphere.

The implications of Equation (7-34) concerning the needed values for Γ_c and $(dT/dr)_c$ have been discussed by Marlborough and Roy (1970), by Castor and Cassinelli (1973), and by Thomas (1973). The rate of mass loss may be determined by means of Equation (7-19). Doing this reduces to finding the value of the product,

$$\dot{M} = 4\pi r^2 \varrho_c a_c. \tag{7-35}$$

Marlborough and Roy (1970) considered the requirements for determining a solution in an *isothermal* atmosphere which fits to a hydrostatic model atmosphere at depth, and in which an outward accelerating force proportional to the force of gravity acts. They write the proportionality factor as $\Gamma(r)$. Marlborough and Roy found that no acceptable solution exists for Equation (7-17) in the isothermal case. Inside the sonic point, Γ must be less than 1 in order for the fit to a hydrostatic model atmosphere to be made, while outside the sonic point, Γ must be greater than 1 in order that the flow be supersonic. A discontinuity in Γ is required at r_c. No regular function can be found for $v(r)$ which will behave as the flow velocity in a wind is believed to behave. One conclusion that they reach is that, if an acceptable solution to Equation (7-17) is to be obtained, the temperature gradient must not be set identically equal to zero.

The first attempt to model a wind which is driven by the force of radiation was made by Lucy and Solomon (1970), who used the concept of an isothermal atmosphere. They also assumed without presenting justification for this step, that the sonic point in their model atmosphere was close to the stellar surface. Work on numerical models for flowing atmospheres which are powered by the force due to radiation absorbed in lines is reported in the section *The Line Spectrum from Moving Three-Dimensional Model Atmospheres.*

The effect of the force of radiation as a driving force in an optically thin atmosphere has been investigated by Castor and Cassinelli (1973). These authors assume that the force due to radiation may be evaluated by means of Equation (7-28), with the value for the line-absorption coefficient put equal to zero. They replace the actual continuous absorption coefficient by a gray value and thus are able to eliminate the integral over frequency. Castor and Cassinelli show that, if one postulates that the rate of mass loss, $\dot{M}$, is like what is deduced for luminous O and B stars, namely $\dot{M} \approx 10^{-6}$ $M_\odot$ yr^{-1}, then an acceptable solution to Equation (7-17) can be found only if Γ_c exceeds about 0.99. This value of Γ_c is found by requiring that the sonic velocity lie in the range from 10 to 20 km s^{-1} i.e, that T_c lie in the range 10000 to 50000 K) and that the density at the critical point be such as to give the observed value for $\dot{M}$. The sonic point must lie relatively close to the stellar surface in order to obtain the needed density in the flow.

Castor and Cassinelli (1973) investigate the conditions which must exist near the sonic point in the type of model atmosphere which they postulate. They conclude that an atmosphere has to be heated if it is to acquire supersonic flow velocities. Unless each element of volume acquires additional heat at the time it acquires additional momentum, supersonic flow will not occur. Castor and Cassinelli suggest that the needed heating may be the result of the absorption of radiation. However, the absorptive properties of O- and B-type stellar atmospheres are not such as to make this a viable proposal.

In passing, it should be noted that, in a real spherical atmosphere, if *isothermal* conditions are to be maintained as postulated by Marlborough and Roy (1970) and by Lucy and Solomon (1970), the stellar atmosphere must acquire heat as the radius increases. This is because the temperature decreases steadily outward in a spherical atmosphere which is in radiative equilibrium (see Figure 7-15), in contrast to its behavior in a planar atmosphere which is in radiative equilibrium. Neither Marlborough and Roy nor Lucy and Solomon suggest a source for the heat which their treatments imply is acquired by the flowing atmosphere.

The conditions which are required so that supersonic flow may occur in stellar atmospheres of all types have been investigated independently by Thomas (1973). Thomas is interested in determining what the observation of mass loss implies about the physical conditions in deep layers of the star. Thomas shows that the greatest depth at which mass loss can have a significant effect on the structure of the model star lies well above the base of the model photosphere. Thomas discusses the implications of Equation (7-17) for the flow velocity, and he shows that, if one wishes to ensure that the sonic point occurs in layers in which the density is relatively large (in order that $\dot{M}$ may be relatively large), one must either increase the term $a^2(r)$ $(1 - d \ln T/d \ln r^2)$ or decrease the term $GM(1 - \Gamma)/r$ (or both at appropriate relative rates). The first option corresponds to putting energy into the flow from some nonradiative source, while the second corresponds to putting momentum into the flow. The papers on the effects of the force of radiation reviewed above discuss the latter option.

Thomas (1973) shows that the arguments followed by Lucy and Solomon (1970) and by Castor and Cassinelli (1973) are equivalent to using the density distribution which is valid for hydrostatic equilibrium and considering the effects of radiation pressure as an accelerating force on material arranged according to this density distribution. He argues that the Lucy-Solomon limit on the rate of mass loss (that which can be obtained by the absorption of radiation in one resonance line) is a lower limit to the rate of mass loss to be expected from a hot star, while that estimated by Castor and Cassinelli is a strong upper limit on what may be obtained as the result of the action of radiation pressure in the case that there is no heating from nonradiative sources. Thomas continues his analysis to show that, if one permits a very small outward velocity v_0 (in our notation) to be imposed on the subphotospheric layers, then this is sufficient to drive the observed winds from stars.

A key factor in the picture suggested by Thomas (1973) for understanding stellar winds is the existence of a field of nonthermal kinetic energy below the photosphere. It is postulated that this source of energy provides the small outward increments of velocity, $v_0(r)$, which feed energy, by means of shocks, into the flow to heat the flow and cause the outflow of mass at the observed rate. The results of the presence of an initial photospheric motion, $v_0(r)$, are what Thomas works out. He implies that the presence of such a storage mode for nonradiative energy in a star is plausible, but he does not specify any particular cases. The major statement made by Thomas is that the *known existence* of mass flow from stars implies the existence of a source of nonradiative energy in the subphotospheric layers of the star. Thomas chooses to describe this source of energy as a field of kinetic energy, which could be due to pulsation or convection in the subphotospheric region. The nonradiative kinetic energy postulated by Thomas is supposed to be like the thermal energy of the particles in the following sense. A field of motions exists, but it does not imply drift motion except in a stochastic sense. The theory is presented in mathematical detail in Cannon and Thomas (1977), and a stability analysis is carried out there.

According to Thomas (1973) and Cannon and Thomas (1977), a small flow in the subphotospheric regions of a star may be sufficient to provide for the heating seen in the mantle of the star, and it will provide the mass flux which is observed. However, factors other than subphotospheric motion may be important for causing the conditions observed to exist in the mantle. In this context, it is important to determine the mechanism which causes superheating in the mantle. Such superheating is seen in almost all types of stars.

Parker (1981) has used the formal solution of the time-dependent hydrodynamic equations which describe conditions in the photosphere, corona, and wind of the Sun to investigate how blocking upward flow from the photosphere of the Sun will change the wind which arises because of the presence of a hot corona. The change is undetectably small. Parker points out that to obtain conditions such as those known to exist in the corona of the Sun requires the deposit of much more energy than can be available in any subphotospheric flow in the Sun just large enough to account for the known rate of mass loss from the Sun. Consequently, Parker claims that the conclusion of Thomas (1973) that small subphotospheric adiabatic flows will generate coronas and winds for dwarf stars like the Sun cannot be supported.

In the case of stars for which gravitational binding just above the photosphere is small because of the action of radiation pressure and/or low gravity, the situation is complex. In this case, the problem cannot readily be treated by analytical methods.

Since there is much evidence for the presence of superheated plasma and moderately dense winds in the mantles of O and Wolf-Rayet stars, it seems probable that the actions of the mechanisms which cause the heating and propulsion have much more to do with determining the density of the wind than does possible subsonic flow in deep layers of the star. It seems unlikely that a subphotospheric flow at subsonic speeds, as proposed by Thomas (1973), would be able to furnish the amount of heating and outward-directed motion which is seen to occur in the mantles of the hot stars.

Hearn (1975a, 1975b, 1983) has studied the model winds from hot stars which may result from the presence of a small hot corona just above the photosphere. Hearn emphasizes that it may be possible to understand the origin of mass loss from O and B stars by finding an appropriate density and temperature for a hot corona. The corona cannot be *very* hot, because then O, Wolf-Rayet, and B stars would be intense sources of X rays, and they are not. Lucy and Solomon (1970) were among the first to show that a temperature of more than 10^7 K throughout the stellar wind would be required if the observed winds from O and B stars were to be generated by the process which is believed to generate and propel at least the slow component of the solar wind. Very high temperatures in large volumes of the mantles of O and B stars are excluded. However, very high temperatures may occur in small parcels of the mantles of O stars because X rays are observed from some O stars. Theory to explain the generation of X rays is reported in the section *Atmospheric Stability: X Rays.*

Hearn (1975a, 1975b) investigated the particular problem of whether the winds from OB supergiants can be the result of the presence of a corona with T_e near 10^6 K. He investigated several sources for the heating. Hearn found that the existence of a small hot corona is not a truly satisfactory mechanism for creating winds from O and Wolf-Rayet stars. In particular, he showed that radiatively driven sound waves cannot deposit enough energy in the postulated coronal layers to keep a wind of the observed density flowing. If one postulates the presence of a corona, one has to find an adequate source of heating for the corona. Hearn used the device of a "minimum-flux" corona to select particular solutions from the many that are possible. An important result of Hearn's work is the attention which he draws to the fact that the rate of mass loss from a star and the conditions in the mantle of the star are as much determined by the energy gains and losses which occur in the mantle as by the momentum-transfer processes which occur there.

Hearn and Vardavas (1981a, 1981b) and Vardavas and Hearn (1981) have developed numerical models of the coronas which may be generated around O9 supergiants as a result of the deposition of a specified amount of mechanical energy deep in the photosphere. They set up a self-consistent method for finding the physical state of the atmosphere, and they note that the equations which define the problem must be solved with conditions imposed at the interface with the interstellar medium, as well as at the interface with a photosphere which is in hydrostatic equilibrium. It is assumed that the only forces acting on the gas are the attractive force of gravity and the dispersive force due to the gas pressure at the ambient temperature in the corona. No contribution of momentum from wave pressure or radiation pressure is included. Thus, the models of Hearn and Vardavas illustrate the formation of a wind when energy alone is added to the atmosphere. These calculations describe one limiting case of what may occur in nature.

Four models have been derived by Hearn and Vardavas for the formation of a mantle around a supergiant O9 star which is emitting radiation at a rate of 5.24×10^{13} erg cm^{-2} s^{-1} from its surface. We shall briefly report here the results when a flux of mechanical energy of 10^4 erg cm^{-2} s^{-1} is added in the photosphere and when a flux of 10^6 erg cm^{-2} s^{-1} is added.

In the first case, an extended corona is formed in which the maximum temperature is 8.90×10^5 K and the sonic point lies at 9.5 stellar radii. The maximum temperature occurs at about 1.3 stellar radii. Mass is lost from this model star at a rate of $3.64 \times 10^{-14} M_{\odot}$ yr^{-1}. In the second case, only a small corona is developed. It has a thickness of 0.03 stellar radii, and it lies between 1.033 and 1.065 stellar radii. The maximum temperature reached is 6.98×10^5 K, and there is no mass loss because the critical point lies at several hundred stellar radii from the star.

These model mantles do not have properties like those we believe to exist for the mantles of O-type supergiants. Quite clearly, the observed winds from O stars are formed as a result of the deposition of momentum as well as energy in the interface layers between the photosphere and the mantle.

The implications of the presence of a small hot corona in the mantle of an O or Wolf-Rayet star for the infrared spectrum from the star are reviewed in the following section.

H. Radio and Infrared Flux from Spherical Atmospheres: Mass Loss

Any distant large volume of fully ionized gas having a cosmic composition, a density in the range 10^4 to 10^9 cm^{-3}, and an electron temperature of a few times 10^4 K will emit free-free radiation at a rate which can be detected in the radio and infrared wavelength ranges. Typical astronomical sources include the H II regions associated with O and Wolf-Rayet stars, as well as some point sources which have been identified with particular O and Wolf-Rayet stars. The purpose of the theory, which will be reviewed in this section, is to provide an interpretation of the observed radio and infrared fluxes from O and Wolf-Rayet stars. Information such as the density in the emitting volume as a function of the distance from the center of the star, the electron temperature, and the size of the emitting region is sought. A specific mechanism explaining how the observed plasma reached its observed physical state underlies each theory which has been proposed for interpreting the radio and infrared fluxes from stars.

Existing theories for explaining the radio and infrared emission from astronomical sources fall into two general classes. The first class contains theories which interpret the observations of extended sources such as the H II regions and planetary nebulae enveloping some O and Wolf-Rayet stars. The second class contains theories which interpret point sources confined to small spherically symmetric volumes around stars. The linear radii of these volumes of plasma are sufficiently small that the cross section of the emitting region is not resolved by the telescope and detector.

In theories of class 1, it is usually assumed that the particle density in the emitting volume can be described by giving the constant density or by giving the relative fractions of the emitting volume which are filled by parcels of plasma having the one or the other of two densities. In the case of theories of class 2, it is usually assumed that the particle density decreases as r^{-n}, where n is a small positive number of the order of 2. The appropriate value of n is determined by the mechanism which is assumed tc cause the plasma to exist around the star at a detectable density. Often the electron temperature, T, is assumed to be constant in the emitting volume. Olnon (1975) has calculated the radio spectra to be expected from some particular distributions of the emitting gas.

In the theories of class 2 (see, for instance, Wright and Barlow, 1975, and Panagia and Felli, 1975), it is usually assumed that the plasma is the result of the uniform, spherically symmetric outflow of gas with each flow line directed along a radius. The flow velocity is assumed to be constant; usually it is put equal to the largest outflow velocity which is seen by means of absorption lines in the stellar spectrum. However, this assumed mechanism for the formation of the radiating plasma may be questioned.

The quiet Sun radiates a radio flux at 6 cm of approximately 10^6 Jy. This radio flux is known to come from the transition region between the photosphere and the corona. If one supposes that the observed rate of mass loss from the Sun, about $10^{-14} M_\odot \text{ yr}^{-1}$, is linked to the radio emission from the Sun in the manner which is assumed to be true for O and Wolf-Rayet stars, one will expect to detect from the Sun a radio flux at 6 cm of only 10^3 Jy. The observed flux is about 10^3 times larger than this because a much denser plasma exists in the transition region of the Sun than is implied by the observed flow of gas from the Sun. The reason why a moderately dense hot plasma exists around the Sun is not fully understood at this time. However, it is believed that magnetohydrodynamic events, as well as the deposition of heat and momentum from mechanical waves, are important for determining the density, volume, and physical state of this plasma.

It seems premature, when discussing O and Wolf-Rayet stars, to make a priori the hypothesis that magnetohydrodynamic events and the deposition of heat and momentum from mechanical waves play no significant role in establishing the density and volume of the hot plasma detected around O and Wolf-Rayet stars. Any suspended hot plasma will emit radio and infrared radiation, just as the hot plasma does in the transition region between the solar photosphere and corona. If the measured radio flux from O and Wolf-Rayet stars comes from a similar suspended hot plasma, at least in part, then the rates of mass loss from O and Wolf-Rayet stars may be less than the rates which are deduced by means of the theories of Wright and Barlow (1975) and of Panagia and Felli (1975). This question has been explored by Underhill (1983a, 1984a). An important question to which we shall return is whether or not there are any observational facts which can guide us toward making an accurate estimate of the fraction of the observed radio and infrared emission from O and Wolf-Rayet stars which is related to the rate of mass loss in the manner which is assumed to be true by Wright and Barlow and by Panagia and Felli for *all* of the observed radio flux.

The plan for this section is to review the theories of Wright and Barlow and of Panagia and Felli for radio flux and then the theories of the infrared excess emission. Finally, a brief discussion of the state of our theoretical knowledge and modeling procedures will be given. The results of applications of the existing theories are summarized in the section *Estimated Rate of Mass Loss from O and Wolf-Rayet Stars*.

1. Radio Spectrum. Wright and Barlow (1975) made the following assumptions so that they could provide a relationship between the observed radio flux and the rate of mass loss from a star:

1. The emitting region consists of spherical shells of fully ionized gas in which the density of ions follows the relation $n(r) = Ar^{-2}$. Here A is a constant which must be assigned. In the case of outflow at a constant velocity v, the constant A

is determined by the rate of mass loss from the star, $\dot{M}$. Then

$$A = \dot{M}/4\pi \, \mu m_H v, \qquad (7\text{-}36)$$

where μm_H is the weight of a typical ion in the plasma. In most of the applications to be found in the literature, the gas is assumed to have solar composition. Then μ is of the order of 1.2. If the gas is assumed to be composed chiefly of helium which is fully ionized, $\mu \rightarrow 4$. This case is sometimes considered for Wolf-Rayet stars.

2. The density of electrons, n_e, is γ times the density of ions, $n_e = \gamma n_i$. In the case of a fully ionized plasma having solar composition, γ is a number of the order of 1.1. When the plasma is composed chiefly of doubly ionized helium, $\gamma \rightarrow 2$.

3. The average charge per ion in the plasma is Z. In the case used in applications to O stars, Z is taken to be 1.1. For fully ionized helium, $Z \rightarrow 2$.

4. The temperature, T, is assumed to be constant throughout the emitting plasma.

5. Radiation at radio frequencies is emitted as a result of free-free encounters between electrons and ions. The source function is assumed to be the Planck function for the temperature T. Since, in the radio wavelength range, we are concerned with wavelengths of several cm while T may be of the order of 20000 K, one may write $B_\nu(T) = 2kT\nu^2/c^2$.

6. The optical depth per unit path length in the plasma is the result of absorption due to free-free interactions in the fully ionized plasma, corrected for stimulated emission. From Allen (1973), p. 102, one finds that the absorption coefficient corrected for stimulated emission is

$$\kappa'_{ff} = 3.692 \times 10^8 \, (1 - \exp(-h\nu/kT)) \, Z^2 g_\nu(T) \, T^{-1/2} \nu^{-3} n_e n_i. \qquad (7\text{-}37)$$

Here $g_\nu(T)$ is the Gaunt factor at frequency ν and temperature T; it can be evaluated from the formula given by Allen on p. 103. For radio wavelengths, one may use the approximate form,

$$\kappa'_{ff} = 1.98 \times 10^{-23} Z^2 \, g_\nu(T) T^{-3/2} \lambda^2 n_e n_i, \qquad (7\text{-}38)$$

with λ expressed in cm.

The radio flux which is measured by a distant observer is that flux which emerges from all parts of the sphere of ionized plasma which are not optically thick at the frequency ν. The distance of the observer is D, measured in some linear unit. Wright and Barlow (1975) solve the relevant problem of radiative transfer under the conditions listed above and find that the rate of mass loss in $M_\odot$ yr^{-1} may be found from the relation,

$$\dot{M}/v = 0.095 \, \mu Z^{-1} \, (\gamma g_\nu)^{-1/2} \, \nu^{-1/2} \, D^{3/2} \, S_\nu^{\,3/4}. \qquad (7\text{-}39)$$

Here the velocity of outflow, v, is in km s^{-1}, the distance between the star and the observer is in kpc, and the radio flux, S_ν, is in Jy.

An expression for the effective radius, $R(\nu)$, of the region outside of which the radio flux is emitted has been derived by Wright and Barlow. It is

$$R(\nu) = 2.8 \times 10^{28} \, (\gamma g_\nu)^{1/3} \, (Z/\mu\nu)^{2/3} \, T^{-1/2} \, (\dot{M}/v)^{2/3} \mathrm{cm}. \qquad (7\text{-}40)$$

The optical depth along a radius from infinity to this point is 0.244.

If one substitutes the appropriate values of the constants describing the plasma of cosmic composition (μ, Z, and γ) and the value of g_ν

for a wavelength of 6 cm and a temperature of 22000 K (which is a reasonable temperature for the radio emitting region), one finds that

$$\log R(6\ \mathrm{cm})/R_\odot = 9.229 + 0.667 \log (\dot{M}/v). \tag{7-41}$$

Equation (7-41) is useful for estimating the approximate inner radius of the region emitting radio flux in the immediate neighborhood of a star which is losing mass. In it $\dot{M}$ is expressed in $M_\odot$ yr^{-1} and v is in km s^{-1}. If T is different from 22000 K and the slight dependence of g_ν on T is ignored, then R(6 cm) will change as $(22000/T)^{1/2}$. For Equation (7-41) to be valid, $\dot{M}/v$ must be large enough that an optical depth of 0.244 is reached at 6 cm when $R(\nu) >> R_*$.

For an O-type supergiant, $\dot{M}/v$ may be about 10^{-9} in the selected units, while the radius of the star may be about 36 $R_\odot$. Then R(6 cm) will be about 47 R_* or 1.2×10^{14} cm. For a Wolf-Rayet star, it is sometimes suggested that $\dot{M}/v$ may be about 10^{-8}, while the radius of the photosphere may be about 10 $R_\odot$. Then R(6 cm) will be about 790 R_* or 5.5×10^{13} cm.

We see from Equation (7-36) that, when $\dot{M}/v$ is large, the density of the plasma near the star will be large, with the consequence that the radio flux observed by a distant observer will come chiefly from an outer shell which is rather far from the star.

The value of $\dot{M}/v$ used in the present example for a Wolf-Rayet star is probably near the upper limit of what may actually occur. If, as in the case of the Sun, part of the radio flux from a Wolf-Rayet star arises from gyroresonance radiation in a plasma which is not flowing from the star in the spherically symmetric manner assumed by Wright and Barlow (1975), then the value of $\dot{M}/v$ estimated by means of Equation (7-39) will be reduced, and $R(\nu)$ will be smaller than in the example given here (Underhill, 1984a).

Panagia and Felli (1975) have gone through similar arguments to those followed by Wright and Barlow. They use a representation for the free-free absorption coefficient which explicitly includes an expression for the Gaunt factor, valid in the radio range. They derive an equation for the relationship between the rate of mass loss, the constant velocity of outflow, and the flux, S_ν, at a radio frequency ν. When $\dot{M}$ is expressed in units of $M_\odot$ yr^{-1} and v is in km s^{-1}, the result of Panagia and Felli becomes

$$\dot{M}/v = 2.94 \times 10^{-9} Z^{-0.5} \left[\frac{\mu}{1.2}\right] \left[\frac{\nu}{10\ \mathrm{GHz}}\right]^{-0.45} \left[\frac{T}{10^4\ \mathrm{K}}\right]^{-0.75} \left[\frac{D}{\mathrm{kpc}}\right]^{1.5} \left[\frac{S_\nu}{\mathrm{mJy}}\right]^{0.75} \tag{7-42}$$

Here the units for each variable are indicated in the square brackets. It is assumed that $\gamma = 1.0$. Panagia and Felli also derive an expression for a representative radius for the emitting region. As a representative radius they adopt that radius within which half the emission is produced. Their result is similar to that of Wright and Barlow, and it can be expressed as

$$\log R(\nu)/R_\odot = 9.283 - \log K + 0.667 \log (\dot{M}/v), \tag{7-43}$$

where

$$\log K = 0.7 \log \nu + 0.45 \log T + 0.667 \log \mu - 0.333 \log Z. \tag{7-44}$$

Here ν, T, and μ are to be expressed in the units indicated in Equation (7-42). In Equation (7-43) $\dot{M}$ is in $M_\odot$ yr^{-1} and v is in km s^{-1}.

For the case of radiation at 6 cm, $T = 22000$ K, $\mu = 1.2$, and $Z = 1.1$; log K is -0.070. The constant term in Equation (7-43) then becomes 9.353. This means that the representative radius of Panagia and Felli is 1.33 times that of Wright and Barlow.

Panagia and Felli also investigated what the shape of the radio spectrum would be if the density in the emitting region followed a density law of the form $N_e(r) \sim r^{-n}$, with n taking values between 1.5 and 3.5. A case with n

< 2 could be obtained if the material were suspended in large magnetic loops or if there was flow at a constant velocity, but the degree of ionization increased outward; one with $n >$ 2 would imply that electrons were being progressively removed from the plasma at large distances (e.g., by recombination and/or by condensation of the material to form grains as the gas cooled). The spectrum which is obtained is a power law in λ, $S_\nu \sim \lambda^{-\alpha}$, with n and α taking the following related values:

n	1.5	2.0	2.5	3.0	3.5
α	-0.10	0.60	0.95	1.16	1.30

A fully ionized plasma at a constant temperature and flowing at constant velocity from a star results in a spectral index $\alpha = 0.60$ in the radio wavelength range. Values of α larger than 0.6 are observed for the radio region in the case of Wolf-Rayet stars (see, for instance, Barlow, 1979, and Barlow et al., 1981). When the slope of the free-free emission spectrum from infrared to radio wavelengths is considered, α is usually found to be near 0.76. These facts imply that $n > 2$. This conclusion may be interpreted to mean that the degree of ionization in the plasma is decreasing outward, or that the terminal velocity has not been reached in the region in which the infrared excess is formed. Barlow (1979) has considered the possibility that, in the volume observed to be emitting at radio frequencies, the gas has cooled adiabatically by a significant amount from its temperature in the volume observed to be emitting at infrared wavelengths and has dismissed it as being of negligible importance.

If the picture underlying the theory of Panagia and Felli (1975) is correct, then the fractional uncertainty in the value of $\dot{M}/v$ derived by means of Equation (7-42) can be found from

$$\left\langle \frac{\Delta(\dot{M}/v)}{(\dot{M}/v)} \right\rangle = 5^{-1/2} \left[\left(\frac{\Delta\mu}{\mu}\right)^2 + 0.25 \left(\frac{\Delta Z}{Z}\right)^2 + 0.5625 \left(\frac{\Delta T}{T}\right)^2 + 0.5625 \left(\frac{\Delta S_\nu}{S_\nu}\right)^2 + 0.4773\,(\Delta \mathrm{Mod})^2 \right]^{1/2} \tag{7-45}$$

Here we have replaced the fractional uncertainty in the distance, $\Delta D/D$, by its equivalent uncertainty in the distance modulus, Mod, of the star, expressed in magnitudes, and we have treated the individual errors as though they add quadratically. There are uncertainties in the appropriate values of Z, μ, and T to use for the plasma because we are not completely sure of the mean charge on each ion, the composition of the gas, nor of its temperature.

A typical value for the fractional error in $\dot{M}/v$ may be estimated by noting that the error in measuring the radio flux and the errors in the factors describing the radiating plasma may have the values $\Delta S_\nu/S_\nu = \Delta Z/Z = \Delta T/T = \Delta\mu/\mu = 0.3$, while ΔMod may be 0.5 mag. In this case, the root mean square fractional uncertainty in $\dot{M}/v$ is 0.25.

2. Infrared Spectrum. The first interpretations of the infrared free-free excess emission from Wolf-Rayet stars (Hackwell et al., 1974; Cohen et al., 1975) use the very simple model of a sphere of ionized plasma around the star. In the plasma, the electron density, n_e, the ion density, n_i, and the temperature have constant values, independent of radius. The radiation is emitted as a result of free-free encounters between the ions and the electrons.

If the emitting volume is optically thin at a wavelength λ (in μm), then it is straightforward to deduce from the emissivity of the plasma, which is $\varkappa_\nu' B_\nu(T)$, that the flux received from a volume of radius R_S cm will be

$$F_\lambda(ff,\ \text{opt. thin}) = 7.17 \times 10^{-16} Z^2 \lambda^{-2} g_\nu T^{-1/2} \exp(-c_2/\lambda T) D^{-2} n_e n_i R_S^3 \,. \tag{7-46}$$

The units of F_λ are Wm^{-2} μm^{-1}. In Equation (7-46), g_ν is the Gaunt factor; it is approximately equal to unity in the infrared wavelength region. The distance of the star is D kpc. The particle densities are expressed per cm^3.

If the distance of the star is known, knowledge of the observed infrared excess fluxes in the wavelength region where the plasma is optically thin coupled with an estimated value for the temperature of the plasma will lead to an estimate of the effective radius of the emitting region.

Because free-free opacity decreases toward short wavelengths at a rate varying between λ^2 and λ^3 (see Equation (7-37)), the emitting plasma may be expected to become optically thin at relatively short infrared wavelengths. Consequently, if one plots log λF_λ against log λ, using the observed amount of infrared excess radiation determined by subtracting an estimate for the flux from the star from the observed total flux at an infrared wavelength, the points will form a curve which shows a maximum. Let the wavelength at which the peak value of log λF_λ occurs be λ_p. Then from Equation (7-46), one finds that

$$\lambda_p T = 14400. \qquad (7\text{-}47)$$

Here λ_p is measured in μm and T in kelvins. Use of Equation (7-47) permits one to estimate T in the plasma which is emitting the infrared free-free radiation.

In the ionized plasmas around stars, one may expect the particle density to decrease outward. This will certainly be true if a wind is flowing from the star. Consequently, a radius, R_S, determined from Equation (7-46) will probably be a lower limit to the size of the emitting region. The degree to which it is reasonable to assume that the temperature is constant in the plasma can only be determined after one has postulated a model for the heating and cooling processes that go on in the plasma. In most applications of this theory, T is assumed to be constant.

In the case of winds which are expanding uniformly from the star, one may interpret the free-free infrared excess radiation of O and Wolf-Rayet stars by means of theory such as that developed by Wright and Barlow (1975) and by Panagia and Felli (1975). The chief modification required is to use the full expression for $B_\nu(T)$ in place of the Rayleigh-Jeans approximation for those infrared wavelengths where T and λ are of such a size that $c_2/\lambda T$ is not "small". In the case that the Rayleigh-Jeans approximation is valid, the equations derived by Wright and Barlow may be used after substituting an appropriate value for g_ν. The formulas of Panagia and Felli include explicitly an expression for g_ν. Since this expression is valid only in the radio domain, their formulas require modification if they are to be used with infrared fluxes for estimating $\dot{M}/v$.

Barlow and Cohen (1977) have measured the infrared flux from a few O stars and some B-type supergiants, and they have investigated what the observed infrared excesses might mean in terms of the rate of mass loss from the star. Application of the expressions of Wright and Barlow (1975) for $\dot{M}/v$ and $R(\nu)$, with the temperature in the emitting plasma being put equal to the assigned effective temperature for the star (which falls in the range from 34500 K at type O9 V to 50000 K at type O4f), leads to inconsistent results when excess fluxes at $\lambda \leqslant 10\ \mu$m are used. Barlow and Cohen conclude that the flow must be accelerating in the part of the mantle from which the infrared spectrum comes. This may be so, but it should be borne in mind that the expressions of Wright and Barlow (Equations (7-39) and (7-40)) are based on using the Rayleigh-Jeans approximation for the Planck function. This expression overestimates the value for the Planck function when λ is not sufficiently large. This simple fact may account in part for what Barlow and Cohen found.

To obtain typical values of $\dot{M}/v$ for each star from its infrared excess, Barlow and Cohen adopted the velocity law which they had deduced by trial and error for P Cygni from the observed infrared and radio excesses of P Cygni. Because P Cygni is a very unusual star (see, for instance, Underhill and Doazan, 1982, Chapter 4), one may doubt that the velocity law adopted by Barlow and Cohen is appropriate for O Stars. It is very difficult to estimate the level of uncertainty in the values of $\dot{M}/v$ determined by Barlow and Cohen from the infrared excesses at $\lambda \leqslant 10\ \mu$m. There is uncertainty in the numerical results not only because of uncertainty about the shape of the velocity law and the adopted temperature in the plasma, but also because of uncertainty whether or not the concept of uniform spherical outflow is correct for fixing the density pattern of the parts of the mantles of O stars in which the infrared spectrum is formed.

A theory for interpreting the spectra of O and Wolf-Rayet stars in the spectral range from 0.5 to 10 μm has been developed by Cassinelli and Hartmann (1977); they have computed some theoretical spectra in support of their ideas. The theory has been applied (Hartmann and Cassinelli, 1977) to interpreting the infrared spectrum of the Wolf-Rayet star, HD 50896, and to interpreting the differences between the eclipse curves of HD 193576 = V444 Cygni in the visible and infrared ranges (Hartmann, 1978a). In addition, Hartmann (1978b) has expanded the simple theory to be valid for the case of a thin disk surrounding a star. He has made some estimates of what the presence of a disk would imply concerning the polarization of the light from a Wolf-Rayet star.

Cassinelli and Hartmann (1977) consider the infrared spectrum that will be produced by a spherical plasma surrounding a hot star. The hot star has a photosphere which can be represented by a classical model atmosphere. The spectrum radiated by the star may be found accurately enough at long wavelengths by finding the flux which will emerge from a model atmosphere in which the temperature law follows a simple approximation for the case of gray opacity. Cassinelli and Hartmann postulate that the density in the spherical plasma around the star varies as r^{-n} and that the temperature in this part of the model varies as r^{-m}. The opacity in the plasma is postulated to be due to free-free encounters between ions and electrons. By following arguments similar to those of Wright and Barlow (1975), Cassinelli and Hartmann conclude that, at a long wavelength, the luminosity radiated from the model can be written as

$$L_\lambda = 4\pi R_\lambda^{\ 2} \ . \ \pi B_\lambda(t_\lambda = 1/3). \qquad (7\text{-}48)$$

Here R_λ is the radius where the plasma becomes opaque, and this, they claim, may be put equal to the radius where t_λ is 1/3. (Wright and Barlow (1975) assumed the effective radius to be that where t_λ is 0.244.) The Planck function is to be evaluated at the temperature corresponding to the radius where t_λ is 1/3. The monochromatic optical depth, t_λ, is found by integrating along a radius from infinity (the position of the observer) to R_λ. Thus R_λ is determined by solving the equation,

$$\int_{R_\lambda}^{\infty} \varkappa_\lambda(ff)\ \varrho \mathrm{d}r = 1/3, \qquad (7\text{-}49)$$

and making use of the fact that $\varkappa_\lambda(ff)$ varies as $\varrho^2\lambda^2T^{-3/2}$. The Gaunt factor is put identically equal to unity. This approximation is rough, but usable in the infrared; it is unsatisfactory at radio wavelengths, as Wright and Barlow, and Panagia and Felli (1975) have noted.

Cassinelli and Hartmann (1977) substitute the *Rayleigh-Jeans approximation* for $B_\lambda(T)$ with the result that they can write

$$\lambda L_\lambda = 4\pi^2 R_\lambda^{\ 2} \ . \ 2c^2kT(R_\lambda)\lambda^{-1}. \qquad (7\text{-}50)$$

After substituting for R_λ and $T(R_\lambda)$ and performing some algebra, they attain the result,

$$\lambda L_\lambda \sim \lambda^{-s}, \qquad (7\text{-}51)$$

where

$$s = \frac{6n - \frac{5}{2}m - 7}{2n - \frac{3}{2}m - 1} \qquad (7\text{-}52)$$

(We have changed the sign of s from the definition of Cassinelli and Hartmann, who wrote $\lambda L_\lambda \sim \lambda^{+s}$, because we wish to emphasize the relationship between s and the power-law exponent α of Panagia and Felli. With our definition of s, we find that the α of Panagia and Felli is equivalent to s–1.)

The following matrix of values relates n, m, s, and α when the Gaunt factor is unity:

n	1.5	2	3	3	3
m	0	any	0	−1	−3
s	1	1.67	2.2	2.08	1.95
α	0	0.67	1.2	1.08	0.95

Panagia and Felli considered the case of a constant temperature, $m = 0$, and they allowed for the dependence of the Gaunt factor on ν and T which is valid in the radio range. Cassinelli and Hartmann note that acceptable solutions require that $|m|$ should lie in the neighborhood of 1 and that, when n is very large and therefore the density gradient is steep, the shape of the spectrum should approach that of the Rayleigh-Jeans approximation for the Planck function.

An important point to note about Equations (7-51) and (7-52) is that they are valid *only* for those wavelength-temperature regimes for which the Rayleigh-Jeans approximation for the Planck function is valid. This means that $c_2/\lambda T$ must be smaller than about 0.10, which means that λ must be longer than λ_{RJ}, where λ_{RJ} depends on temperature as shown in Table 7-5.

In the case of O and Wolf-Rayet stars, very few observations have been made at wavelengths longer than 4 μm. Most infrared observations are confined to the range shortward of the wavelength of the L filter. Some observational results showing the shape of the infrared-excess flux spectra of Wolf-Rayet stars are presented in Underhill (1980a); others for O stars are given by Castor and Simon (1983). A plot of log λF_λ against log λ results in a curve with a turnover at λ_p, followed by a linear section which indicates that at long wavelengths $\lambda F_\lambda \sim \lambda^{-s}$.

The theory of Cassinelli and Hartmann (1977) for interpreting the observed dependence of $\lambda F_\lambda(ff)$ on λ in the *observed* spectral range is valid only if the electron temperatures in the emitting plasma are greater than 36000 K. There is little direct evidence about the true value of the electron temperature at distances from the photospheres of O and Wolf-Rayet stars where the infrared excess is believed to originate.

To interpret the shape of the infrared-excess energy spectrum at wavelengths between 1 and 10 μm, one should use the full expression for the Planck function, just as is done when writing Equation (7-46) and deriving condition (7-47). The spectrum should not be expected to follow a power law in λ from the very long wavelengths of the radio range, through the far-infrared range to the near infrared. Because of the form of Equation (7-46), a turnover is expected when one plots log λF_λ against log λ; the position of the maximum in this curve gives information primarily about the temperature in the emitting plasma.

The conclusions of Hartmann and Cassinelli (1977)—that Wolf-Rayet mantles are optically thick in the wavelength region shortward of 5 μm and that the turnover which is seen in a plot of log λF_λ against log λ reflects the effects of a rapid drop in the density at the edge of the emitting region, or shell—result from their use

Table 7-5
Minimum Wavelengths for the Validity of the Rayleigh-Jeans Approximation

T (K)	λ_{RJ} (μm)	T (K)	λ_{RJ} (μm)	T (K)	λ_{RJ} (μm)
8000	18	24000	6	72000	2
14400	10	36000	4	144000	1
18000	8	48000	3		

of the Rayleigh-Jeans approximation in wavelength and temperature regimes for which it is not valid. Consequently, their deductions concerning the density profiles and velocity laws valid in the mantles of Wolf-Rayet stars have no firm basis in reality.

The applications by Hartmann (1978a, 1978b) of the theory of Cassinelli and Hartmann (1977) to the Wolf-Rayet binary HD 193576 = V444 Cygni and to Be stars are likewise of doubtful validity because Hartmann is chiefly concerned with plasmas in which the temperatures are too low to justify use of the Rayleigh-Jeans formula in the wavelength region of the observations.

Cassinelli and Hartmann (1977) developed numerical models to explore the shape of the infrared spectrum to be expected from a plasma around a hot star in the case that a hot coronal region ($T \sim 2 \times 10^6$ or 2×10^5 K) of relatively small extent was present close to the photosphere of the star. They found that the spectrum would have a shape more or less as predicted by the simple theory with the addition of a hump in the energy distribution between 10 and 100 μm. The precise size of the hump and its shape depend on the exact details of the assumed temperature and density laws. The shape of the spectrum shortward of 10 μm is little changed by the presence of the hot gas from its shape when only a plasma having a temperature equal to the effective temperature of the star is present. The numerical results of Cassinelli and Hartmann at wavelengths shortward of 10 μm are biased by their use of the Rayleigh-Jeans approximation for the Planck function, but the general trend of their results is probably indicative of what may occur.

The infrared excesses to be expected from O stars have been predicted by Castor and Simon (1983) and by Groot and Thé (1983) using detailed model atmospheres. Both studies adopt a parameterized velocity law in order to model the effects of an accelerating wind, and each assumes that mass is conserved in spherical shells. Castor and Simon assume that the electron temperature in the infrared-emitting region is equal to the effective temperature of the star, while Groot and Thé assume that the electron temperature is equal to 0.75 T_{eff} in one set of models and equal to 2×10^5 and 2×10^6 K in other models. Both studies imply that a source of energy exists in the mantle. This is because T_e will decrease outwards in a spherical atmosphere when no energy in addition to that carried by the radiation field is present (see Figure 7-15). Both groups conclude that the rate of mass loss from an O-type star can be determined reliably from a measured infrared excess only if the shape of the velocity law near the star has been determined accurately by the analysis of other information.

3. Discussion. The theoretical interpretations of the radio and infrared fluxes from O and Wolf-Rayet stars are based on very simple models for the density and temperature structure of the mantle. We have seen that the theories for interpreting the radio emission assume a constant temperature. On the other hand, some numerical experiments have been made to explore how the presence of a temperature rise, followed after a short distance by a temperature drop, will change the shape of the infrared spectrum. No significant effect was found in the range of wavelength in which infrared observations have been made. The infrared flux is increased a little at long wavelengths. The effects of more realistic model temperature laws can be explored once a physical theory for the heating and cooling processes in the mantle has been developed to establish the shape of realistic temperature laws.

Free-free emission in rather low-density gas around the star appears to be a principal source of the excess long-wavelength radiation observed for O and Wolf-Rayet stars. This flux of energy arises as a result of encounters between the ions and electrons in the mantle. A hot star shows an infrared excess and emission at radio wavelengths only if it is surrounded by a mantle in which the density of charged particles is greater than what would be expected to be present were the density gradient in the outermost parts of the atmosphere solely that valid for layers of gas in hydrostatic equilibrium. The factors affecting the intensity of the radiation are the mean charge on the ions, Z, the mean atomic weight of the gas, μ, thus its composition, and the relative numbers of ions and electrons in each element of volume in the mantle, γ.

Assuming spherical symmetry, a composition, and power laws involving radial distance for the variation of particle density and temperature is one way of describing the model mantle. Another way is to assume that the plasma has the properties of a "cosmic" plasma which is fully ionized and to adopt a constant temperature. Either of these procedures fixes μ, Z, and γ. Then one needs only to adopt a density law to complete the definition of the properties of the emitting plasma. The formulas which permit one to estimate $\dot{M}/v$ from an observed radio flux (see Equations (7-39) and (7-42)) are based on the concept of flow at a constant velocity and use of the Rayleigh-Jeans approximation for the Planck function. It is possible to use an explicit velocity law giving the outflow velocity as a function of radius and derive a value for the infrared and radio flux emitted as a function of the adopted value of $\dot{M}$. This approach for finding $\dot{M}$ from a radio flux has not generally been taken, although it has been used (Barlow and Cohen, 1977; Tanzi et al., 1981) to turn observed infrared excesses into estimates for $\dot{M}/v$. This procedure is similar (in principle) to the procedures to be discussed in the section *Rate of Mass Loss from Hot Stars*. The cautionary remarks of Castor and Simon (1983) and of Groot and Thé (1983) should be heeded.

If one is to feel secure about the values of $\dot{M}$ found from the observed radio and infrared fluxes, one must be able to establish two pieces of information by means of other observations: (1) the proper geometrical structure to represent the parts of the mantle which are emitting the infrared and radio flux, and (2) the velocity law followed by the material which is escaping from the star.

Because the light from some O and Wolf-Rayet stars is intrinsically polarized, one may conclude that the mantles of some O and Wolf-Rayet stars are not spherically symmetric. We mentioned in the introduction to this section that, in the case of the Sun, most of the radio flux at 10 cm does not come from the solar wind; it comes from gas trapped in the transition region between the chromosphere and corona of the Sun. The structure of this part of the Sun is strongly influenced by the structure of the magnetic field which emerges from the solar photosphere. Likewise with O and Wolf-Rayet stars, we should recognize that a structured magnetic field may be present. The average field strength may be below detectable levels with present observing equipment, but the effects of the presence of a structured magnetic field on the density and velocity structure of the mantles of O and Wolf-Rayet stars should not be ignored when one is considering possible models for the density, temperature, and velocity structure of the mantles of O and Wolf-Rayet stars. If some plasma is being more or less suspended in the immediate neighborhood of the photospheres of O and Wolf-Rayet stars by the action of magnetic fields, then the values for the rate of mass loss deduced by means of the theories reviewed in this section will be upper limits to what is actually occurring (see Underhill, 1983a, 1984a).

The difficult question yet to be answered is how to find observational criteria which will permit us to differentiate between the two possible types of models for the structure of mantles: (1) spherical outflow throughout the mantle, and (2) a suspended plasma near the photosphere, surrounded by a halo of plasma which is flowing from the star. Obviously, before one can answer this question, one must understand what forces are initiating outflow and accelerating it, as well as what mechanisms are heating and cooling the gas. Possible answers to this question are described in Chapter 8, which reviews the full set of information now available about O and Wolf-Rayet stars relative to the existing theoretical descriptions for the atmospheres of these types of stars.

IV. THE LINE SPECTRUM FROM MOVING THREE-DIMENSIONAL MODEL ATMOSPHERES

In this section, we shall summarize: (1) the chief points of the theories used for interpreting the parts of the line spectra of single O, Of, and Wolf-Rayet stars which are formed in the mantle of the star, and (2) the chief results attained. More than 100 years ago, Abney (1877) proposed interpreting the broad lines of some B stars in terms of rotation of the star. Since that time, quite a few theoretical studies have been carried out for the purpose of predicting the shapes of the lines formed in moving extended atmospheres. A review has been presented in Underhill and Doazan (1982), Chapter 6, of work of this type relevant to early-type stars. A summary of the radiative-transfer theory which is used may be found in Mihalas (1978), Chapter 14. Van Blerkom (1973) has summarized the theory available in 1971 for interpreting the line spectra of Wolf-Rayet stars. Little advance on that specific problem has been made since then, even though the theoretical ideas required are similar to those needed to interpret O-type spectra.

All of the published predicted profiles of lines formed in winds have been obtained using ad hoc model atmospheres in which analytical expressions, or tables of values, are used for $T(r)$, $v(r)$, and $\varrho(r)$. In each case, a fixed rate of mass loss from the star, $\dot{M}$, is postulated, and the density distribution is found from the adopted velocity law by using Equation (7-19). Spherical symmetry is assumed.

In the first studies of lines formed in moving atmospheres (see, for instance, Beals, 1929; Menzel, 1929; and Rublev, 1961), the emphasis was on deducing likely forms for $T(r)$, $v(r)$, and $\varrho(r)$. Some applications in the case of Wolf-Rayet stars are given by Rublev (1963) and by Lyong (1967). Later, after it was realized that mass loss was a common phenomenon for luminous early-type stars, the emphasis shifted to deducing values of $\dot{M}$ from the observed shapes and strengths of the absorption lines formed in the expanding mantle (see, for instance, Conti and Garmany, 1980; Garmany et al., 1981; Gathier et al., 1981; Olson, 1981; and Olson and Castor, 1981). In the case of Wolf-Rayet stars, interest has been concentrated on finding the temperatures of the photosphere and the mantle, and the composition. We shall try to show the major trends of the theoretical developments which have taken place in the theory of line formation in moving atmospheres which do not rotate.

Theory which is aimed at understanding the line spectra of quasars is often applicable also to the spectra of O and Wolf-Rayet stars. In addition, some interesting cases of line formation in moving atmospheres have been studied with T Tau stars in mind.

A. Concepts Used

The basic geometric picture which is adopted for all of the studies to be reported here is that of line formation in a spherical mantle surrounding a spherical photosphere. In most cases, the photosphere of the star is represented as a spherical core which radiates a continuous spectrum. In the neighborhood of the wavelength of the line under study, this continuous spectrum is put equal to the Planck function, $B_\nu(T_*)$. Here T_* is a temperature which is assigned to the star; it is assumed to be about equal to the effective temperature of the star. In a few studies, the method used to find the spectrum permits the investigator to take account of line formation in the photosphere of the star as well as in the mantle. In all cases, possible continuous opacity in the model mantle, due to electron scattering or other sources, is ignored.

The conservation laws should be the source of the relationships which constrain the ranges of temperature, density, and velocity which will be encountered in a moving extended model mantle. In practice, in the papers which present predicted line profiles from moving extended model mantles (see Table 7-6), the conservation laws are ignored except for the statement that mass is conserved in spherical shells (Equation (7-19)).

Arbitrary expressions are used for $T(r)$ and $v(r)$ in each case. Then use is made of Equation (7-19) to find $\varrho(r)$. An important factor which is rarely worked out in detail is the changing degree of ionization, as r increases, of the ion under investigation. In many of the studies, the electron temperature in the mantle, T_e, is assumed to be constant, and it is frequently set to a value between 0.8 and 0.95 T_*. Because $T(r)$ decreases outward in a static spherical model atmosphere which is in radiative equilibrium (see Figure 7-15), the assumption of a constant T_e implies that energy is being deposited in the moving model mantle at values of r larger than R_*, the photospheric radius.

The following expression is frequently used for a velocity law:

$$v(r)/v_\infty = (b + (1 - b)(1 - R_*/r))^\beta. \quad (7\text{-}53)$$

This expression implies that

$$dv/dr = (1 - b)\beta R_* r^{-2} (b + (1 - b)(1 - R_*/r))^{\beta-1} v_\infty. \quad (7\text{-}54)$$

Here b is a constant which is assigned according to the choice of the investigator, v_∞ is the terminal velocity in the wind (i.e., the value of $v(r)$ approached as $r \to \infty$), and β is an assigned constant. Frequently, the value 1/2 is used for β, and β is sometimes set to 0 or to 1.0. When β is positive, $v(r)$ increases outward, which means an accelerating wind. When β is negative, dv/dr is negative, and there is a decelerating wind. In this case, the flow velocity is represented as decreasing from a large outward-directed value, v_0, at $r \simeq R_*$ toward the limiting value v_∞ at large r. The value v_0 is acquired by some unexplained means.

Any model mantle made in the foregoing manner is ad hoc. The actual velocity and temperature laws in a mantle should satisfy Equations (7-17) and (7-23) when appropriate boundary conditions are used. However, the state of knowledge, at the time the work to be reported was done, was insufficient for a rigorous attack to be made on the problem of modeling winds. Before the temperature and velocity laws which have been used are accepted as being true for particular stars, they must be confirmed by analyzing information other than the profiles of resonance lines.

B. Radiative Transfer in Moving Atmospheres

The problem of radiative transfer in a moving mantle is difficult mathematically, and it is solved in an approximate manner, particularly in the first studies. In most cases, because resonance lines are studied, a two-level model atom is used. The earliest work considered the case of line formation by isotropic coherent scattering. In later papers, the case of line formation by isotropic scattering with complete frequency redistribution is considered. In some studies, collisional de-excitation of the upper level of the line is taken into account.

When model atoms/ions consisting of many levels are considered, the distribution of the model atom/ion over its several energy states is found by solving the equations of statistical equilibrium. Frequently, the degree of ionization of the relevant atom/ion is postulated to remain constant at all values of r, or to follow a prescribed function of r. This is often true when the investigators study a two-level atom. The number density of the atom/ion of interest is found from $\varrho(r)$ by specifying a particular composition for the model mantle; sometimes a degree of ionization is also specified.

The hypothesis of Sobolev (1947, 1958) that the velocity v_D corresponding to the Doppler width of the line is negligibly small in comparison to the flow velocity at all points in the atmosphere (more accurately, relative to the component of flow velocity along all lines of sight from points in the extended mantle to the observer) is used in most studies. This hypothesis is sometimes called the narrow-line hypothesis or the supersonic approximation.

Radiative transfer in moving *planar* model atmospheres has been studied by Kalkofen (1970) and by Noerdlinger and Rybicki (1974).

They have demonstrated that the source function is not very sensitive to the details of the assumed velocity law, and that outflow will produce a shortward displaced, asymmetric absorption trough. In the case of planar atmospheres, an outward temperature rise is needed to obtain an emission component.

The equations describing the transfer of radiation in a moving *three-dimensional* atmosphere can be found in Rybicki (1970), together with a lucid discussion of the implications of the hypotheses introduced by Sobolev (1947, 1958). Castor (1970) has developed a compact way of handling radiative transfer in a spherical model atmosphere when the source function is described in terms of the escape-probability concept introduced by Sobolev. By using the Sobolev approximation, Castor simplifies the integrations necessary to find line profiles and obtains an explicit expression for the line profile.

The problem of radiative transfer in a moving extended atmosphere has been studied also by Lucy (1971). He made the Sobolev narrow-line hypothesis in order to limit the region which must be considered when evaluating the source function at any particular frequency in the line profile, and he used the fluid or comoving frame in his description of the radiation-transfer problem. Lucy derived his final results in terms of moments of the radiation field; this allowed him to obtain explicit expressions for the resulting line profiles.

In both of these studies, as well as in many later studies by other authors (see Table 7-6), it is assumed that $v(r)$ increases outward. Then each line of sight cuts each constant-velocity surface only once. A clear exposition of the significance of constant-velocity surfaces may be found in Hummer (1976). There the basic mathematics of radiative transfer in moving atmospheres is outlined.

When dv/dr is negative, a line of sight may cut a given constant-velocity surface (defined by use of the narrow-line approximation) more than once. Then the value of the source function at one point in the model mantle may be influenced by radiation coming from a distant point. This problem—radiative transfer in a model mantle where $dv/dr < 0$ at some places—has been studied by Kuan and Kuhi (1975), Grachev and Grinen (1975), Marti and Noerdlinger (1977), and Rybicki and Hummer (1978). These papers, particularly the last two, should be consulted for details concerning this complex problem. On the whole, the local value of the source function may be little affected by radiation from distant points. However, the local value for the force due to radiation is sensitive to the shape of the constant-velocity surfaces and to the interlocking of radiation that may occur.

Use of the fluid or comoving frame to solve the problem of radiative transfer in expanding model atmospheres has been strongly advocated by Mihalas and his colleagues. When a powerful enough computer is available that large matrices of data may be manipulated, it is not necessary to make the narrow-line hypothesis of Sobolev (1947, 1958), and the true shape of the line-absorption coefficient may be considered. Explicit directions for finding the source function in model atmospheres where $dv/dr \geqslant 0$ everywhere have been given by Mihalas and Kunasz (1978). They are applied by Kunasz (1980) to obtain an understanding of the He II spectrum of Of stars. Mihalas and Kunasz are able to study many-level model atoms having a level structure like that of hydrogen; they apply the principle of statistical equilibrium to find the variation of the degree of ionization and excitation for the model atom as a function of radius. Mihalas and Kunasz predict line profiles for subordinate lines of a hydrogen-like spectrum, and they give information about the resonance lines. This numerical method for solving the problems of radiative transfer in a moving extended atmosphere also allows one to take account of radiative transfer in an underlying photosphere (see Kunasz, 1980).

The method of Mihalas and his colleagues has been refined by Hamann (1981), who has applied it to the study of resonance lines in spherical expanding mantles. Hamann is able to show that use of the narrow-line hypothesis to simplify determining the source function in the specified problem, as done, for instance, by Castor (1970) and by Lucy (1971), does not introduce serious differences in the source function from the results of a more exact solution obtained using the true value for v_D at all times. However, quite important differences arise in the results from the part of the problem concerned with integrating the emergent specific intensity from the model mantle to find the line profile which will be seen by a distant observer. Removing the approximation introduced into this part of the radiative-transfer problem by the escape-probability method (Castor, 1970), brings about significant changes in all parts of the line profile.

In all of the studies of spectrum formation in a moving extended spherical atmosphere reviewed here, the inner boundary of the model mantle is coincident with the edge of the photosphere of the star. An outer boundary is arbitrarily established according to some reasonable hypothesis. Generally, r_{max} lies in the range 10 to 50 R_*.

In Table 7-6 information is given about the form for $v(r)$ used, the type of model atom studied, and how the degree of ionization of the atom/ion under study is assumed to change as a function of radius. In the cases indicated by the letter T, the degree of ionization is found by solving the equations of statistical equilibrium, allowing for photoionization and recombination. In none of these cases is ionization and recombination by means of states with two excited electrons taken into consideration. Allowing for doubly excited states can modify the results of a statistical equilibrium analysis by a significant amount (Bhatia and Underhill, 1986). Sometimes (cases noted by F) an explicit expression is used for the dependence of the degree of ionization on radius or on the outflow velocity (see, for instance, the work of Olson, 1978, 1981; of Castor and Lamers, 1979; and of Hamann, 1981). In other cases, the degree of ionization of the ion under study is not explicitly mentioned; these cases are denoted by C. An asterisk is placed in the last column of Table 7-6 to indicate those studies in which grids of line profiles are presented to illustrate the effects on the profiles of the resonance and subordinate lines of different choices for the parameters defining the adopted functions representing $v(r)$ and the fractional abundance of the ion of interest. In the other papers, only a few illustrative profiles are presented.

Among the studies listed in Table 7-6, only those of Bertout (1977) and Surdej (1979) deal explicitly with infalling material. The others deal with the outflow of material from a central source. Comparison of the results for outflow with deceleration with those for outflow with acceleration shows that the major differences in the shape of the resulting line profiles occur near the undisplaced position of the line. In the case of decelerating outflow, the drop in intensity from the peak of the emission to the deepest point of the shortward displaced absorption trough is very steep, and the deepest point of the absorption trough lies close to the undisplaced position of the line. In the case of accelerating outflow, the descent from the peak of the emission component to the deepest point of the absorption trough is less steep. In addition, the deepest point of the absorption trough moves toward a wavelength displaced shortward from the center of the line by an amount of the order of $\lambda - \lambda_0 = v_\infty \lambda_0 / c$ (see Figure 7-17), where some typical predicted profiles are displayed.

The theoretical "wind" profiles shown in Figure 7-17 are from Hamann (1981). We use them to illustrate two points. The first is how the line profile is changed from what results when the narrow-line approximation of Sobolev (1947, 1958) is made and the problem of radiative transfer is solved by the escape probability method of Castor (1970) and when one solves the problem of radiative transfer accurately using the comoving frame and the assumption that v_D / v_∞ is 0.1. The second is to illustrate how the line profile changes when the distribution of ions in the model mantle is changed.

Hamann assumes that the ratio of the Doppler velocity, v_D, to the terminal velocity of outflow; v_∞, is constant through the model

Table 7-6
Major Characteristics of Studies of Line Formation in Moving Atmospheres[a]

Reference	Velocity Law	Model Atom	Degree of Ionization	Line Width	Grid of Profiles
Rublev, 1961	O,A	2	C	N	
Rublev, 1964a	O,D	2	C	N	
Castor, 1970	O,A	2	F	N	
Lucy, 1971	O,A	2	T	N	
Grachev and Grinen, 1975	O,D	2	C	N	*
Kuan and Kuhi, 1975	O,A; O,D	M	T	N	
Oegerle and van Blerkom, 1976b	O,A	M	T	N	
Marti and Noerdlinger, 1977	O,D	2	C	N	
Bertout, 1977	I,D	2	T	N	
Mihalas and Kunasz, 1978	O,A	M	T	W	
Olson, 1978	O,A	2	F	N	*
Rybicki and Hummer, 1978	O,D	2	C	N	
Castor and Lamers, 1979	O,A	2	F	N	*
Surdej, 1979	O,A; I,A	2	C	N	*
Surdej, 1979	O,D; I,D	2	C	N	*
Kunasz, 1980	O,A	M	T	W	*
Rumpl, 1980	O,A	2	C	N	*
Hamann, 1981	O,A	2	F	W	*
Olson, 1981	O,A	2[b]	C	B	*

[a]Abbreviations used:
O — outflow—material in the mantle moving away from the surface of the star.
I — inflow—material in the mantle moving toward the surface of the star.
A — acceleration outward, $dv/dr > 0$ at $r > R_*$.
D — deceleration outward, $dv/dr < 0$ at $r > R_*$.
N — narrow-line approximation of Sobolev.
W — allowance is made for the Doppler width of the line.
2 — two-level model atom.
M — many-level model atom.
C — dependence of the degree of ionization on r is not explicitly stated.
F — degree of ionization follows a specified function of r.
T — degree of ionization (and excitation) calculated from the adopted $T(r)$ and the radiation field.

[b]Only two levels of the atom are explicitly considered—the ground level and an excited level from which the subordinate line under study arises. Atoms reach the excited level as a result of absorbing radiation in an optically thick line.

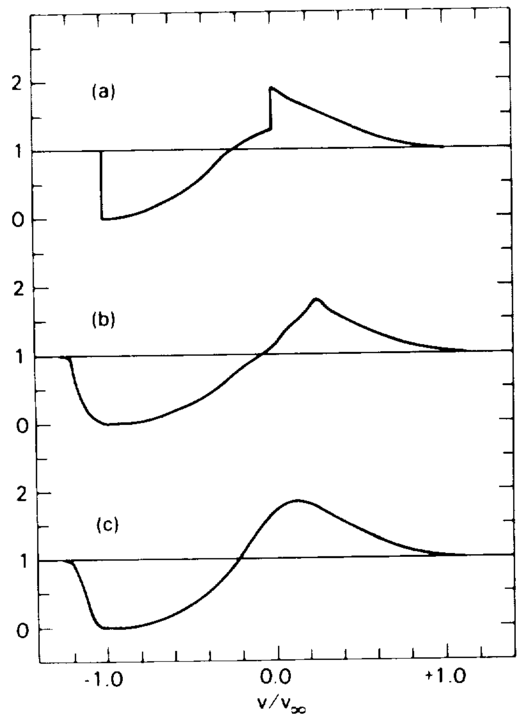


Figure 7-17. Theoretical "wind" profiles from Hamann (1981). Profile (a) is for the case of the Sobolev (1947, 1958) approximation with $\alpha = 0$, $\beta = 1/2$. Profiles (b) and (c) are calculated using the comoving frame and assuming that v_D/v_∞ is 0.1. In case (b), $\alpha = 0$, $\beta = 1/2$; in case (c), $\alpha = +2$, $\beta = 1/2$.

mantle. Here v_D represents the broadening of spectral lines by thermal motion and by microturbulence. Line formation is postulated to be the result of isotropic scattering with complete redistribution. No account is taken of continuous sources of opacity in the mantle. The photosphere provides an amount of energy, I_c, at the wavelength of the line. The quantity I_c is independent of angle and frequency.

The velocity law used by Hamann has the form,

$$v(r)/v_\infty = (1 - b/r)^\beta. \qquad (7\text{-}55)$$

The constant b is fixed by the condition that $v(r) = 0.01\ v_\infty$ when r is equal to R_*. The variation of the line-absorption coefficient with radius (essentially the number density of absorbers as a function of radius) is assumed to follow the relation,

$$\varkappa(r)/\varkappa_0 = r^2(v(r)/v_\infty)^{-1}q(r). \qquad (7\text{-}56)$$

Here $q(r)$ represents the depth dependence of the degree of ionization of the line-forming ion. Hamann explores the case of

$$q(r) = (v(r)/v_\infty)^\alpha. \qquad (7\text{-}57)$$

The various cases treated by Hamann are specified by the values of the constants $\varkappa_0$, α, and β. The profiles shown in Figure 7-17 are all for a strong line, $\varkappa_0 = 10$, and a velocity law with $\beta = 1/2$.

In the middle panel of Figure 7-17, the case for $\alpha = 0$, or constant degree of ionization, is illustrated. In the bottom panel, $\alpha = +2$, which represents the case when the degree of ionization decreases inward.

By combining Equations (7-55), (7-56), and (7-57), one may obtain an expression for the variation of the number density of absorbing ions as a function of radius. The variation of the numbers of absorbing ions in our selected cases are displayed in Figure 7-18. The case $\alpha = 0$, $\beta = 1/2$ corresponds to a large density of absorbers near the photosphere, the point where $\log r/R_* = 0$, followed by a decrease outward which eventually varies as r^{-2}. The case with $\alpha = +2$, $\beta = 1/2$ corresponds to the presence of very few absorbers at the photosphere, followed by a rapid increase in number in a region close to the photosphere. After a short distance outward, the number of absorbers begins to decrease, and eventually, far from the photosphere, it decreases as r^{-2}.

Let us first discuss the effects of using an improved treatment of the problem of radiative transfer in a moving atmosphere and of allowing for the Doppler width of the line to correspond to motions of the order of $0.1\ v_\infty$. The changes which occur can be seen by comparing panels (a) and (b) of Figure 7-17. Two things happen: (1) the abrupt rises at $v/v_\infty = -1$ and 0 of panel (a) disappear when the transfer problem is handled in detail, and (2) the

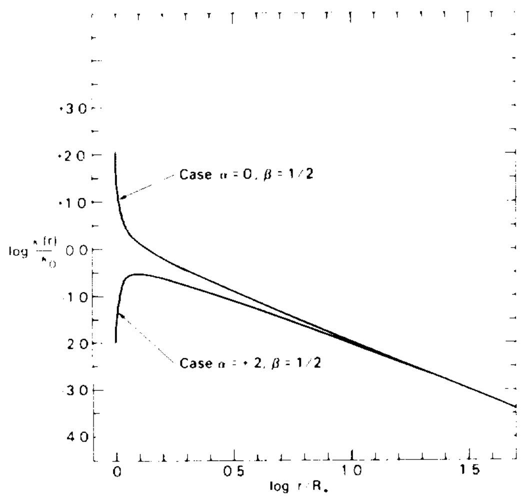


Figure 7-18. The variation of $x(r)/x_0$ with radius for two of the cases calculated by Hamann (1981). Here $x(r)$ is the value of the line-absorption coefficient at the center of the line. The constant x_0 is set to 10 for the profiles shown in Figure 7-17.

steepness of the shortward edge of the absorption trough and the steepness of the rise to peak emission are both softened when v_D/v_∞ differs from zero by a significant amount. In both panel (a) and panel (b), the area of the emission component is nearly as large as the area of the absorption trough. The peak emission reaches an intensity between 1.5 and 2.0 times the level of the continuum.

In panel (a), the crossover from absorption to emission occurs close to the wavelength where v/v_∞ is -0.24; in panel (b), it occurs near -0.08. The profile of panel (b) is what one may expect for a strong line when microturbulence is large. The absorption trough is saturated, remaining close to zero in the range from $v/v_\infty = -1$ to about -0.85; the rise to the continuum at the shortward end of the profile is somewhat sloping. The presence of a gradual incline at the shortward end of the profile is a signature for the presence of a field of motion which may be described, more or less, by the term "microturbulence". The slow slope from peak emission to the deepest point in the absorption trough is typical of what occurs in the case of an accelerating atmosphere with the number of scatterers per cm^{-3} decreasing outward.

Profiles for resonance lines have been predicted using the narrow-line approximation by Olson (1978), Castor and Lamers (1979), and Surdej (1979) for a number of velocity laws and values of the total opacity in the line center. In all cases when the line opacity is large, the profile has a shape like that shown in panel (a) of Figure 7-17. Different assumptions concerning the velocity law and the variation of the degree of ionization through the atmosphere result in minor modifications of the part of the line profile between the deepest point of the absorption and the peak of the emission. When the total number of scattering ions along the line of sight becomes small, the absorption does not dip to zero intensity. Then the deepest point moves toward the undisplaced position of the line. Also, the emission reaches undetectably low values before reaching a wavelength corresponding to $v/v_\infty = +1$.

Whenever the process of line formation is represented as being due to isotropic scattering of an underlying continuum of radiation, the equivalent width of the emission component is equal to that of the absorption component to a good approximation. Castor and Lamers (1979) have investigated what happens when an underlying photospheric absorption component is present. Then the emission component is reduced in strength, and the absorption component is strengthened, particularly in the wavelength range occupied by the photospheric line.

Comparison of panels (b) and (c) of Figure 7-17 shows the differences in the profiles which result when the scattering ions are concentrated close to the photosphere (panel (b)) and when they are concentrated in a region near $r/R_* = 1.6$ (panel (c)). The most significant changes in the profile occur between the points, $v/v_\infty = \pm 0.4$. The relevant distribution functions for the scattering ions are shown in Figure 7-18. The crossover point from absorption to emission shifts shortward from being close to the undisplaced position of the line, and the emission component becomes broader and stronger as the concentration point for the scatterers moves away from the photosphere. These two changes are reminiscent of the differences seen between the profiles of resonance lines in O and Wolf-Rayet stars (see, for instance, Underhill, 1983a).

Castor (1970) calculated some theoretical line profiles which look like the profiles of some lines seen in Wolf-Rayet stars. He obtained typical Wolf-Rayet line profiles only when he introduced a particle-density function for the Wolf-Rayet model mantle that had the general pattern of the curve shown in Figure 7-18 for the case $\alpha = +2, \beta = 1/2$. To obtain a dominant emission component, as seen for Wolf-Rayet stars, it is necessary for the concentration of scattering ions to be very small near the photosphere and to reach a maximum somewhere between $r/R_* \approx 1.2$ and 3.0, or so. Use of a density function such that the density of the scattering ions decreases steadily outward from a large value near the photosphere, as for the case $\alpha = 0, \beta = 1/2$ in Figure 7-18, produces a strong absorption trough with an emission component which has an equivalent width of the order of or slightly smaller than the equivalent width of the absorption trough.

In the case of a line profile formed in a moving atmosphere, it is impossible to correlate the intensity of light at a particular wavelength in the line profile with conditions at a specific radius in the atmosphere. This is because the Eddington-Barbier relationship fails in a moving atmosphere. Light at a given velocity displacement in the line profile comes from the neighborhood of all points along the constant-velocity surface corresponding to the line-of-sight velocity. Constant-velocity surfaces cut across the spherical surfaces of equal density and equal component of velocity directed along a radius. This fact makes it very difficult to deduce precise information about the properties of the moving mantle from the observed shapes of the lines which are formed in the mantle.

In the case of Wolf-Rayet stars, the energy contained in the emission component of a line may exceed that removed by the absorption trough by a factor of 2 to 10. This means that the mantle is adding energy in line wavelengths to the stellar spectrum from a store of energy which is different from the energy carried by the continuous radiation field of the star. The amount of energy provided may be described by means of an emission measure for the particular ionic spectrum under consideration. This subject is discussed in the next section.

Castor and Lamers (1979) have investigated what sort of line profile will result if the Planck function in the mantle has a value greater than that in the photosphere (i.e., if T_e(mantle) > T_e(photosphere)). They find that resonance lines formed in the flowing mantle (wind) may then have strong emission components with saturated absorption troughs. The amount of energy appearing in the emission component becomes large when $B_\nu(T)$ in the mantle is made much larger than the adopted $I_c = B_\nu(T_*)$ for the continuum radiation from the photosphere. An alternative way of expressing this result is to say that postulating a high electron temperature in the mantle results in a large emission measure for the resonance lines formed in the mantle.

Mihalas and Kunasz (1978) and Kunasz (1980) have investigated how the shapes of the subordinate lines of a hydrogen-like spectrum depend on the assumed $T(r)$ and $v(r)$ in the mantle. Kunasz is able to show that the observed fact that, in the spectra of some Of stars, the He II line at 4686 Å (3 - 4) has a strong emission component while the He II line at 3203 Å (3 - 5) has weak or absent emission can be understood in terms of appropriately adjusted temperature and velocity laws.

Olson (1981) has studied the line profiles to be expected for subordinate lines which arise from a level which is in detailed balance with the ground level. These calculated line profiles have shapes like those found earlier by several workers for resonance lines, but in general, the subordinate lines are not saturated. Each absorption trough is accompanied by an emission component which has an equivalent width approximately equal to the equivalent width of the absorption trough. Neither the absorption troughs nor the emission components of these subordinate lines extend to the full width corresponding to an outflow velocity equal to v_∞. This is because the density structure of the model mantles studied is so arranged that excited ions capable of forming the subordinate lines under consideration are found only relatively near the photosphere where the outflow velocity is less than v_∞.

That quite strong resonance lines of N V and O VI are seen in absorption in the spectra of O stars is a fact that is glossed over in the theories of line-driven winds reported above. In most cases, the model winds are assumed to be in radiative equilibrium or to have a constant temperature about equal to the effective temperature of the star. The observed strengths of the N V and O VI lines imply the presence of N^{+4} and O^{+5} ions in such abundances that electron temperatures of the order of a few times 10^5 K are required if the N^{+4} and O^{+5} ions are to be formed by electron collisions. Cassinelli et al. (1978) and Cassinelli and Olson (1979) have noted that the required number of high ions can be formed by X-ray-caused Auger ionization from lower ions which may be in abundance in the mantle when the electron temperature is of the order of the effective temperature of the star. These authors postulate that sufficient X rays are formed in a slab of very hot material at the base of the wind. No mechanism is suggested, however, for heating some parcels of gas at the base of the wind to temperatures near 10^7 K.

Cassinelli et al. (1981) and Waldron (1984) have worked out the details of this scheme for producing the observed high ions in the needed abundance in the winds of O stars. The results are satisfactory for accounting for the observed fluxes of X rays from O stars when account is taken of the changes in opacity in the wind induced by the X rays. At present, there is no truly satisfactory theory for accounting for the levels of excitation and ionization seen in the winds from O and Wolf-Rayet stars. The problem of the possible energy gains and losses in the winds from hot stars needs to be studied in parallel with the problem of the propulsion of the winds in order to obtain a deepened understanding of what happens.

C. Theories for Interpreting Wolf-Rayet Spectra

The emission-line spectra of Wolf-Rayet stars have long presented a challenge to astrophysicists, and interpretations of Wolf-Rayet emission lines have been developed and exploited independently by several authors. We shall mention here a few of these studies, giving references to those papers which may be expected to be available in libraries of moderate size. In these papers, the interested reader can find references to most of the work that has been published on the subject of interpreting the line and continuous spectra of Wolf-Rayet stars. The work done prior to 1965 is summarized in Underhill (1966), Chapter XIII.

We choose not to give an exhaustive review of the published work on Wolf-Rayet spectra because, for the reasons noted below, we believe Wolf-Rayet spectra to be chiefly the result of the deposition of nonradiative energy and momentum in the mantles of these stars, which is in contrast to the point of view taken in some of the published papers in which it is assumed that the observations can be interpreted in terms of heating and ionization resulting *only* from the passage of radiation. In the early papers, as in all recent papers, no attempt is made to account for the observed outflow of matter; in some recent papers, the models are completely ad hoc. Studies of this type are seriously incomplete and inconsistent because they ignore entirely the important part of the problem which concerns how the gas reached the physical state it is seen to be in and how it is maintained in that state without significant change over intervals of at least 100 years. The available studies treat the interactions between radiation and gas in considerable detail.

In the preceding sections of this chapter, we have shown what sort of a spectrum may be expected from a hot atmosphere which is in radiative equilibrium. We have given the results both for atmospheres in hydrostatic equilibrium and for flowing atmospheres. In particular, we have seen that the initiation of flow at supersonic speeds requires the input of both energy and momentum (see the section on the *Characteristic Features of the Stellar Wind Problem*), and that, if the temperature is to rise in the line-forming regions outside the photosphere, energy must be deposited in the gas. We have presented the results of calculations of the profiles of resonance and subordinate lines in a flowing atmosphere (see the section on *Radiative Transfer in Moving Atmospheres*), and we have noted that, to obtain a predominantly emission line spectrum, as for Wolf-Rayet stars, it is necessary to maintain a significant amount of gas in a large volume around the star. In addition, the density distribution must be such that there are fewer emitting ions in a region contiguous with the photosphere than there are somewhat farther out. A suitable density distribution is shown in Figure 7-18.

It is argued in Underhill (1980a, 1981, 1983b) that the effective temperatures of Wolf-Rayet stars are of the order of 25000 to 30000 K. However, the ions which are seen to be present in abundance in the mantles of Wolf-Rayet stars require more energy for their formation than can be supplied by a radiation field equivalent to the effective temperatures of Wolf-Rayet stars. One must conclude that nonradiative energy is being deposited somewhere in the gas flowing from Wolf-Rayet stars.

The amount of energy radiated in the emission lines from Wolf-Rayet stars may be evaluated in terms of an emission measure for the emitting volume, just as is frequently done for the emission lines from gaseous nebulae and for those from the mantles of late-type luminous stars. Use of the term "emission measure" implies that the radiative energy seen in lines may be interpreted in terms of the density of ions and electrons in the effective volume of the emitting region and the probabilities for recombination and for self-absorption. Evaluating the emission measure by means of observations should result in a quantitative estimate of how much energy has to be supplied to the mantle each second in order to balance the losses by radiation. This number can be used to gain information about sources for the nonradiative energy in the mantle.

Since, in the case of most single Wolf-Rayet stars, changes in the strengths of the emission lines have usually been small or undetected over the interval of time in which the stars have been observed, we may infer that, whatever the source of nonradiative energy is in the mantle of a Wolf-Rayet star, this source behaves in a rather steady manner. The source certainly has done so for the 115 years and more that the best-known Wolf-Rayet stars have been observed.

1. Preliminary Estimates of the Composition in Model Wolf-Rayet Mantles. All published interpretations of the relative strengths of the emission lines in the visible spectra of Wolf-Rayet stars start from the premise that helium is the most abundant constituent. We wish here to determine if this is a good assumption from which to start.

The conclusion that helium is the most abundant element in the atmospheres of Wolf-Rayet stars was reached independently by Rublev (1972a) on the basis of simple ideas about the formation of He II and H I lines in the mantles of Wolf-Rayet stars, and by Smith (1973a, 1973b) as a result of a qualitative interpretation of the apparent relative intensities of consecutive lines of the Pickering series ($4 - n$) of He II. The lines of the Balmer series of hydrogen have almost the same wavelengths as those lines of the Pickering series of He II which have an upper quantum number $n = 2m$, where m is the upper quantum number of a Balmer line. For Hβ, Hγ, Hδ, and Hϵ, m = 4, 5, 6, and 7, respectively.

In order to find the relative abundance of helium to hydrogen, the above authors assumed that the H I and He II lines observed in the visible spectra of Wolf-Rayet stars are due to cascades following recombination. They further assumed that there is no self-absorption and that the underlying continuum radiated by the photosphere is flat. The possible presence of absorption lines from hydrogen in the photosphere is ignored, and it is assumed that the continuous opacity due to the mantle is negligible. It is argued in the case of Rublev, but assumed without investigation by Smith, that conditions are such in the mantle that the populations of the upper levels of the H I and the He II lines are in LTE ratios with respect to the numbers of next higher ions present. With this set of assumptions, the apparent intensities of the He II and He II plus H I blends seen in the visible wavelength range are inverted to give the ratio by number of He to H atoms.

Rublev finds He/H $\sim$ 6 to 11 for the Wolf-Rayet stars HD 191765, 192103, 192163, 192641, and 193077; Smith deduces values for He/H between 1 and 20. In addition, Nugis (1975), using simple model atoms, has solved the equations for statistical equilibrium in the mantle and has analyzed the strength of Hβ, deducing that the ratio He/H lies between 4.2 and 42.5 for the stars studied by Rublev (see Nugis, 1975).

Quite apart from the theoretical problem of determining how the populations of the upper levels of H and He^+ behave relative to the populations of the next higher ions, the accuracy of the resulting He/H ratios depend on the accuracy of the observational quantities which go into the solution. Significant observational difficulties exist in measuring accurate equivalent widths or central intensities for the blended and unblended lines of the Balmer and Pickering series. Furthermore, neglect of the photospheric absorption lines, which must be present in the spectra from all photospheres which have a composition such that $N(\mathrm{H})/N(\mathrm{He}) \geqslant 0.05$, means that the intensities of the blends at the wavelengths of the Pickering lines with even upper quantum number will be systematically underestimated. The above statement about photospheric hydrogen lines is a fact which can easily be demonstrated by calculating the spectra of hot model photospheres having different H/He compositions.

In early studies of Wolf-Rayet spectra (see, for instance, the summaries by Underhill, 1968b; 1968c), it was believed that the particle density in the line-forming region of Wolf-Rayet stars might lie in the range 10^{11} to 10^{14} cm^{-3}. Present knowledge (see below) suggests that 5×10^{11} or less is probably a more representative density than 10^{14} cm^{-3}. Consideration of the fact that subordinate lines of C IV and N V are seen in the visible spectra of Wolf-Rayet stars, presumably as the result of cascades following recombination, suggests that the electron temperatures in the mantles of Wolf-Rayet stars may be of the order of 50000 to 100000 K. At particle densities and electron temperatures like these, how much recombination is to be expected for H^+ and He^{+2}? Should one expect the upper levels of the most frequently studied lines (i.e., those with $n \geqslant 4$ for hydrogen, $n \geqslant 7$ for the He^+) to be depopulated as a result of non-LTE effects or may the level populations relative to the numbers of next higher ions be approximately in the LTE ratios?

First let us consider the radiative recombination coefficient to the sum over l of the levels (n, l) for the case of hydrogen-like atoms/ions. From information given by Glasco and Zirin (1964), one finds that the ratio of the recombination coefficient to a level $2m$ of a hydrogen-like ion with charge Z on the nucleus to the recombination coefficient to a level m of H is:

$$\frac{R(\mathrm{ion},\ Z,\ 2m)}{R(\mathrm{H}^+,\ m)} = \frac{Z^4}{2} \quad . \tag{7-58}$$

In the case of He^{+2}, Z is 2 and this ratio is equal to 8. Thus, for two plasmas with similar particle densities of hydrogen ions and electrons or He^{+2} ions and electrons, radiative recombination occurs eight times more rapidly to the upper level $2m$ of an He^+ ion than it does to the level m of an H atom. This is for the case that the temperature is the same in each plasma. When $N_e \sim 10^{11}$ cm $^{-3}$, three-body recombination may be much more probable than radiative recombination.

Now let us consider the question of level populations relative to numbers of ions. Since the $2m$ level of He^+ lies below the ionization limit of He^+ by the same amount of energy as the m level of H lies below the ionization limit of H, the LTE population ratio $N(He^+, 2m)/N(He^{+2})$ will be equal to the LTE population ratio $N(H, m)/N(H^+)$ if the ratio of the partition functions at the estimated temperature and electron density is the same for H and H^+ as it is for He^+ and He^{+2}. The calculations of Sparks and Fischel (1971) indicate that this condition will not be met at the temperatures and electron densities usually considered for Wolf-Rayet atmospheres. The ratio of partition functions is much larger for H than it is for He^+. Can both the $2m$ level of He^+ and the m level of H be expected to be in LTE with respect to their next higher ion in the mantles of Wolf-Rayet stars? Bhatia and Underhill (1986) have shown that the answer is no.

Griem (1963) has shown that, if the populations of the levels of a hydrogen-like atom or ion are to be within 10 percent of their LTE values with respect to the number of ions in the next higher stage in a plasma in which the temperature is T and the electron density is N_e, then the electron density must be such that

$$N_e \geqslant 7 \times 10^{18} \frac{Z^6}{n^{17/2}} \left(\frac{kT}{E_H}\right)^{1/2} \text{cm}^{-3} . \tag{7-59}$$

Here E_H is the ionization energy of hydrogen. Rewriting this expression, we see that the quantum number of the lowest level having a population within 10 percent of its LTE value is:

$$n_0 \geqslant 81.53 Z^{12/17} N_e^{-2/17} T^{1/17}. \tag{7-60}$$

The result is very slightly dependent on T, more so on N_e. We obtain the values of n_0 shown in Table 7-7 when T takes the noted values. Table 7-7 gives results for H, He II, and for C III which has $Z_{eff} = 3.2$ for its upper levels, as well as for C IV or N V, both of which have $Z_{eff} = 4.0$ for their upper levels.

Clearly in the case of hydrogen, the population of the upper level of Hϵ—the level with $m = 7$—will approach its LTE value only when the electron density exceeds 2×10^{11} cm^{-3}. If the upper level of Hδ—the level with $m = 6$—is to have a population approximating its LTE value, the electron density in the mantle must be greater than 10^{12} cm^{-3}. In the case of He II, when the electron density is less than 2×10^{11} cm^{-3}, only the levels with $n \geqslant 12$ have their LTE populations. When the electron density is of the order of 10^{12} cm^{-3}, level $n = 10$ of He^+ may have approximately its LTE population. The Pickering line originating from this level has a wavelength of 4338 Å. The populations of the upper levels of the Pickering lines at longer wavelengths than 4338 Å will not have their LTE values.

In Table 7-7, we give estimates of n_0 for the C III and C IV (or N V) spectra. All of the numbers are much larger than the upper chief quantum numbers associated with the lines of C III, C IV, and N V seen in Wolf-Rayet spectra. Clearly, the relative populations of the excited levels of the carbon and nitrogen ions in Wolf-Rayet atmospheres may be expected to have non-LTE values. The populations are probably depleted relative to what they would be in LTE, with most of the ions being concentrated in the ground level of the ion.

Calculations of non-LTE model atmospheres in radiative equilibrium (see, for instance, Auer and Mihalas, 1972) indicate that in that case the most likely type of departure from LTE populations is a depopulation of the upper levels of hydrogen and He^+ with a concentration of the atoms/ions in the ground level. Thomas (1949) has indicated that, if the electron temperature in the mantle is higher than the effective temperature of the star, the populations of high levels may be above their LTE values. What actually occurs has been studied by Bhatia and Underhill (1986), using model atoms and solving the equations of statistical equilibrium. The departures of the populations of the levels of H, He, and He^+ from LTE values are severe when realistic values of T and N_e are considered.

When Rublev (1972a) and Smith (1973a, 1973b) estimated the relative abundances of hydrogen and helium, each placed significant

Table 7-7
Chief Quantum Number of the Lowest Level in LTE

Ion	H,$Z=1$			He$^+$,$Z=2$			C^{+2},$Z=3.2$	C^{+3},N^{+4},$Z=4$
T (K)	10^5	5×10^4	3×10^4	10^5	5×10^4	3×10^4	10^5	10^5
0.1[a]	9.6	9.2	8.9	15.6	15.1	14.5	21.7	25.5
0.5	8.8	8.5	8.2	14.4	13.9	13.4	20.0	23.5
1.0	8.1	7.8	7.6	13.2	12.8	12.4	18.5	21.7
2.0	7.5	7.2	7.0	12.3	11.8	11.4	17.1	20.0
5.0	6.7	6.5	6.2	10.9	10.6	10.2	15.3	18.0
10.0	6.2	6.0	5.8	10.1	9.7	9.4	14.1	16.5

[a]Electron density in units of 10^{11} cm^{-3}

weight on the observed strengths of the Pickering and Balmer lines which lie longward of Hϵ, and they assumed that the level populations would be in LTE ratios with respect to the abundance of the next higher ion. The results of Bhatia and Underhill show that the populations of the upper levels of hydrogen may be depleted, while those of He$^+$ may be nearly normal. There is no reason to infer that hydrogen is deficient in the atmospheres of Wolf-Rayet stars (Bhatia and Underhill, 1986).

In addition, it should be remembered that the strengths of the blends due to H and He$^+$ have been systematically underestimated by Rublev (1972a), Smith (1973a, 1973b), and Conti et al. (1983) because the presence of photospheric hydrogen absorption lines has been ignored. We must conclude that the evidence for a true depletion of hydrogen in the atmospheres of Wolf-Rayet stars is very weak. In fact, hydrogen lines are seen in absorption in the spectra of some Wolf-Rayet stars, and this is not believed to be due to the presence of a nearby O or B star. It is wise at all times to bear in mind the strong possibility that the composition of the mantles of Wolf-Rayet stars may be normal.

2. Preliminary Estimates of the Electron Densities in Model Wolf-Rayet Mantles. If the temperatures of Wolf-Rayet mantles are of the order to 50000 to 100000 K, the continuous opacity at wavelengths between 1000 and 7000 Å will be chiefly due to electron scattering. Arguments such as those used by Wright and Barlow (1975) to estimate the effective inner radius of the region of an expanding atmosphere which emits radiation at radio wavelengths (see the section on *Radio and Infrared Flux from Spherical Atmospheres: Mass Loss)* tell us that the apparent edge of the photosphere of a Wolf-Rayet star will occur at the radius at which the continuum optical depth in the visible range, measured along a line of sight from the observer, is about 0.24. Only gas outside this radius contributes strongly to the line spectrum. If we knew the effective geometric length of the mantle, we could estimate the typical electron density in the mantle. It is possible to do this in the case of the eclipsing binary star, V444 Cyg (WN5 + O6).

From solutions of the light curves of V444 Cyg measured at continuum wavelengths, one finds that there is an extended envelope around the star in which the opacity is small and which may be attributed to electron scattering (Kron and Gordon, 1950; Cherepashchuk, 1975). According to Kron and Gordon, the length of the electron scattering region is 9 $R_\odot$; according to Cherepashchuk (see his Figure 4), the length is about 14 $R_\odot$. If, in accordance with solutions of radiative transfer in spherical atmospheres, we say that the optical depth in electron scattering of this low-opacity region is 0.24, then the average electron density is 6.0 $\times$ 10^{11} cm^{-3} if the path length is 9 $R_\odot$, and 3.8

$\times\ 10^{11}\ cm^{-3}$ if it is 14 $R_\odot$. We conclude that typically the electron density in the mantle of the Wolf-Rayet star of V444 Cyg may be about $5 \times 10^{11}\ cm^{-3}$.

That this value is about correct is confirmed by the spectroscopic observations of Munch (1950), who found that the absorption lines of the O star of V444 Cyg were broadened when the O star was eclipsed by the outer part of the atmosphere of the Wolf-Rayet star. The amount of broadening corresponds to an optical depth of no more than 0.5 in electron scattering.

The effect of electron scattering in spherically expanding envelopes on the emission lines formed in these envelopes has been investigated by Auer and van Blerkom (1972). They use the Monte Carlo technique to follow photons as the photons are scattered by electrons. Auer and van Blerkom predict what the shape of an emission line will be in the case of no electron scattering and when the optical depth due to electrons is 1.0. They allow for thermal redistribution, and they explore the effects of two different velocity laws. The most conspicuous result is that the emission line develops a significant wing longward. No emission lines in Wolf-Rayet stars have this shape (see Underhill, 1968b).

Isolated emission lines in Wolf-Rayet spectra are always rather symmetrical, except when they are eroded on their shortward side by self-absorption from the material in the wind. Consequently, we conclude that the optical depth caused by the electrons in the mantles of Wolf-Rayet stars is less than 1.0 by a significant factor. Indeed, we know it must be less than about 0.24, if proper attention is paid to the results of solving the equation of radiative transfer in a spherical atmosphere in which the continuous opacity is not negligible.

3. Analysis of Lines by Means of Modified Nebular Theory. This type of theory has been developed by Rublev (1964b, 1972a, 1972b, 1972c); a summary of the results can be found in Rublev (1975). The intensities of emission lines are interpreted as being due to cascades following recombination. It is postulated that the star consists of a core (photosphere) surrounded by an envelope (mantle) and that the ionization is caused by radiation from the photosphere. The radiation field from the star is represented by a blackbody energy distribution at a temperature, T_*. When analyzing the He II lines, Rublev considers only the radiation shortward of the primary ionization limit of He II at 228 Å, and he postulates that the mantle absorbs all of the radiation which emerges from the photosphere. In the case of WC stars, the radiation field is truncated at the primary ionization limit of C IV, which lies at 192 Å; in the case of WN stars, the truncation occurs at the primary ionization limit of N IV, which lies at 160 Å. Rublev's equations allow him to deduce how much energy is required to provide enough He^{+2} ions that the lines of the Pickering series of He II have their observed values. This analysis fixes the parameter, T_*. The mantle is treated as though it extends to infinity, and no statement is made about the density of material in the mantle except that the mantle is opaque to radiation shortward of 228 Å.

The effects of collisions with electrons are neglected, but allowance is made for self-absorption in the first members of the Pickering series. The electron temperature, T_e, in the recombining plasma is estimated from the relative intensities of the upper members of the Pickering series (Rublev, 1972c) and found to be near 2.5×10^4 K. The relative energy distribution of the continuous spectrum in the visible spectral region is evaluated in terms of a spectrophotometric or color temperature, T_c. Typically, T_c is found to be about 1.8×10^4 K (see Rublev, 1972b). Radiative transfer is treated in terms of the escape probability concept, and it is noted that the populations of the lowest levels of He^+ will not correspond to LTE conditions at any of the three temperatures (T_*, T_e, and T_c) used in this theory to describe conditions in the photosphere and mantle of a Wolf-Rayet star.

It is argued that the ratios of the populations of levels with high n to the numbers of He^{+2} ions approach their LTE values for $T = T_e$.

This, of course, implies that collisions are important for establishing the number of He^+ ions in these levels, or that the density of radiation at the relevant wavelengths for photoexcitation is about equivalent to a blackbody energy density. Rublev indicates that photoionization from the level $n = 2$ of He^+ may be significant in some Wolf-Rayet stars, but that photoionization from levels with $n \geqslant 3$ is not important.

Rublev has carried out full solutions for the stars, HD 192103 and HD 192163, using equivalent widths found from the intensity atlases of Underhill (1959). In addition, he has developed a method for estimating T_* for other Wolf-Rayet stars, using the observed equivalent width of He II 4686 and the value of a parameter which he establishes from his detailed study of HD 192103 and 192163. The values of T_* found by Rublev vary from 7.0×10^4 K for HD 184738 (WC9) to 1.13×10^5 K for HD 193793 (WC6). Rublev (1975) considers these temperatures to be quite close to the effective temperatures of the Wolf-Rayet stars which he studies.

The values of T_* found by Rublev are about three times larger than the effective temperatures which result from the definition of effective temperature in terms of the total radiation field from each square cm of surface (see Underhill, 1983b). They give an estimate of the amount of energy necessary to provide the number of ions needed to account for the observed strengths of the He II lines in the visible spectral range when these lines are interpreted as a recombination spectrum. It is not necessary, however, to assume implicitly, as Rublev does, that the only source of energy in the mantle of a Wolf-Rayet star is the radiation field from the photosphere. Recent observations made from space, particularly those of X rays, permit us the freedom of thinking in terms of energy from nonradiative sources heating the mantle.

The spectrophotometric, or color temperatures, of Rublev are low because they describe the apparent shape of the continuous spectrum in the visible range. It is argued in Underhill (1980a) that this shape is the result of the addition of the true continuum from the photosphere and the free-free radiation from the mantle. The amount of free-free radiation increases toward long wavelengths, which is what causes the values of T_c to be low.

The electron temperatures estimated by Rublev are uncertain, as Rublev points out (Rublev, 1972c), but they are probably real. They most likely measure the electron energies in the postcoronal flow outside the region where most of the nonradiative energy is deposited. It may be only a coincidence, but the T_e values of Rublev are about the same as the T_e values estimated from the turnover point of the infrared excess of Underhill (1980a).

The method of analysis developed by Rublev does not yield information on the number density in the line-forming regions of Wolf-Rayet stars. However, deductions can be made about the relative abundances of the observed species from the relative intensities of lines from these several species. It is a basic hypothesis that sufficient He^+ ions are present to absorb all of the radiation between 228 Å and the truncation point due to the primary ionization limit from an abundant species. Similarly, in the case of the abundant species C^{+3}, N^{+3}, and N^{+4}, enough ions are assumed to be present along the line of sight to absorb all of the photons present in the wavelength range available for ionizing the ion under discussion.

In the model mantles of Rublev and his colleagues, the ions are considered to be arranged in layers with the ions requiring the most energy for their ionization lying closest to the photosphere. Justification for assuming a stratified model is usually sought from the observations: (1) that the emission components of the N V lines at 4603.73 and 4619.98 Å, as well as the blended C IV emission lines at 5806 Å, often are narrower than the emission components of He II or He I lines, and (2) that the shortward displacement of the absorption components associated with the N V lines is usually less than that of absorption components associated with He II or He I lines. Sahade (1980) has discussed the observational material bearing on this point. The true interpretation of the observations is obscure owing to the complexity of the problem of line formation and line blending in expanding atmospheres (see the section *Radiative Transfer in Moving Atmospheres*). In addition, it must be remembered that the Eddington-Barbier relationship is not valid in moving atmospheres.

Nugis (1975) has modified the theoretical treatment of Rublev to allow for electron collisions to have some effect on establishing the populations of excited levels. He has solved the equations of statistical equilibrium for some simple model atoms in the case of model mantles in which T_* has a value near 10^5 K and $T_e = 6.0 \times 10^4$ K. Typically, Nugis finds that, in WN stars, the ratio of the number of nitrogen to helium atoms lies in the range 0.05 to 0.1, while for WC stars, the ratio of the number of carbon to helium atoms lies in the range 0.7 to 0.9. In the case of the WN stars, the ratio C/He is found to lie in the range 0.002 to 0.008. We noted above that Nugis, from an analysis of Hβ, found the ratio He/H to lie in the range 4.2 to 42.5.

All of these results on the compositions of Wolf-Rayet atmospheres are acknowledged to be very uncertain, perhaps by a factor 3, because they are very dependent on the model assumptions. It is assumed in the models considered by Rublev (1975) and by Nugis (1975) that the lines from different ions are formed in separate zones, yet the electron temperature is assumed to be constant (at a high value according to Nugis, but near 2.5×10^4 K according to Rublev) throughout the mantle. All of the energy required to form the ions is assumed to come from the radiation field emerging from the photosphere. The actual radiation fields from Wolf-Rayet stars (Underhill, 1983b) will not generate electron temperatures in the mantle as high as those assumed by Nugis.

4. Analysis of Lines When Collisions Dominate. Castor and van Blerkom (1970) have studied the problem of calculating the emission line intensities resulting from a model mantle in which the level populations are the result of radiative transitions and transitions induced by collisions. They solve the radiative-transfer problem in a moving atmosphere by means of the escape probability method of Castor (1970). The theoretical formulation is applied at one representative radius in the model mantle. The star is considered to consist of a photosphere of radius, R_c, and a mantle. The photosphere radiates a continuous spectrum corresponding to a blackbody at temperature T_c, and it is surrounded by the mantle in which material is flowing away from the star at a constant speed. The intensities of the lines from a grid of model mantles are calculated and compared with observed line intensities. Each model mantle is described by an electron temperature, T_e, and an electron density, N_e, an algorithm for relating the number of ions of a given species to the assumed electron density, a representative radius, R_e, and a velocity of outflow, $v(R_e)$. The best model is selected by comparing the observed pattern of line intensities with several computed patterns of line intensities. The electron density, electron temperature, representative radius, velocity of outflow, and composition of the best fitting model mantle are taken to be representative for the Wolf-Rayet star under study.

The method of analysis consists of first determining the representative parameters which are kept fixed, namely R_c, T_c, R_e, and $v(R_e)$. Then, after a preliminary composition has been estimated, the line intensities are found for trial values of $N_e(R_e)$, a precise composition, and $T_e(R_e)$.

This method of analysis was applied by Castor and van Blerkom (1970) to the He II spectrum of HD 192163 and of HD 191765 (both type WN6). A coarse analysis using some of these concepts has been carried out by van Blerkom and Patton (1972) to estimate the H/He ratio in HD 50896 (WN5), while Castor and Nussbaumer (1972) have used the full method to study the C III spectrum of γ^2 Vel (WC8 + O8). Oegerle and van Blerkom (1976a) have interpreted the He I spectrum of MR119 (WN8) by these methods, and Willis and Wilson (1978) have applied the procedures to find information about the He, C, and N content of HD 50896, 191765, 192103 (WC8), and 192163, as well as to find representative values for the electron densities and temperatures in the mantles of these stars. Additional applications have been made by Smith and Willis (1982, 1983).

By using conditions at a single representative radius as input to the equations of statistical equilibrium, all discussion of possible radial structure in the mantle is avoided. In place of considering the effects of particular functions $T(r)$, $v(r)$, and $\varrho(r)$, or an adopted value for $\dot{M}$, Castor and van Blerkom (1970) treat the mantle as though it has properties which do not vary with radius. Within a sphere of radius R_e, the source function and the optical depth in a line frequency are assumed to have values such that:

$$S_\nu(\text{line})(1 - \exp(-\tau_{\text{line}})) = \text{constant}. \quad (7\text{-}61)$$

Outside the radius R_e, the mantle vanishes so far as solving the equations of radiative transfer and statistical equilibrium are concerned. According to normal considerations for a two-level atom, the source function in the line can be found from the expression,

$$S_\nu(\text{line}) = \frac{2hc}{\lambda^3}\left[\frac{N_j g_i}{N_i g_j} - 1\right]^{-1}. \quad (7\text{-}62)$$

When the escape probability concept is used (see Castor, 1970), the optical depth in the center of the line is given by

$$\tau_{\text{line}} = \frac{\pi e^2}{mc}\, gf\lambda\, \frac{R_e}{v(R_e)} \left(\frac{N_i}{g_i} - \frac{N_j}{g_j}\right). \quad (7\text{-}63)$$

Here we consider level i to be the lower level of the line, level j to be the upper level. The quantities g_i and g_j represent the statistical weights of levels i and j, and N_i and N_j represent the populations of these levels at the representative radius, R_e. The fraction, $R_e/v(R_e)$, is equivalent to the inverse of the typical velocity gradient in the mantle at radius R_e. Castor and van Blerkom (1970) show how the expressions noted as Equations (7-61), (7-62), and (7-63) can be used with the equations of statistical equilibrium to find the intensities of the lines emitted by the model mantle.

Let us now review how the constant parameters which describe the model photosphere and mantle are determined. So far as the parameters R_c and T_c which describe the photosphere are concerned, the procedure is as follows. First an approximate value for T_c, the photospheric temperature, is assumed from some outside information, and then R_c is found by requiring that the visual magnitude of the star—assuming that the star is radiating like a blackbody at temperature T_c—be compatible with an adopted M_V. A typical velocity of outflow, $v(R_e)$, is adopted from the observed displacements of absorption components in the visible spectral range. Determination of the representative radius, R_e, and a typical composition and density requires more lengthy consideration. The procedures which are followed amount to doing a coarse analysis of the spectrum. They are reported by van Blerkom (1973) as well as by Castor and van Blerkom (1970).

Castor and van Blerkom (1970) estimate the full extent of the mantle by considering the strengths of three He II lines which reasonably may be considered to be optically thick, namely He II 4686 (3 - 4), 3203 (3 - 5), and 10124 (4 - 5). They show that use of Equations (7-62) and (7-63) with the assumption expressed in Equation (7-61) leads to the result that the total line energy radiated can be written as

$$I_{ji} = \frac{K}{\lambda^4}\, \frac{(1 - \exp(-\tau_{ij}))}{(N_i/g_i)/(N_j/g_j) - 1}. \quad (7\text{-}64)$$

Here K is a proportionality constant. In addition, it follows that the emergent flux at the center of line $j \rightarrow i$ is

$$F_{ji} = 4\pi^2 R^2 S(R)(1 - \exp(-\tau_{ij})) + F_c. \quad (7\text{-}65)$$

Here τ_{ij} is the optical depth at the line center found from Equation (7-63).

By comparing the observed relative intensities (I_{43}, I_{53}, and I_{54}) of the three optically thick lines of He II mentioned above (for each of them, the exponential term in Equation (7-64) is effectively zero), Castor and van Blerkom (1970) are able to deduce a value of $(N_3/g_3)/(N_4/g_4)$ for each star they study. This value, substituted into Equation (7-62), gives a value for the source function of the line He II 4686. Then Equation (7-65) is used to find the radius, R, of the emitting region for the He II spectrum. A value of F_{43} is found from spectrophotometric observations, and a value of F_c is estimated from the radius and temperature of the core, which have already been determined. In this way, Castor and van Blerkom find that the full extent of the He II emitting region is $R = 70\, R_\odot$ for HD 192163.

Castor and van Blerkom next state that the representative radius characterizing the mantle as a whole should be the mean of R_c and R. They adopt $R_e = 40\, R_\odot$ for further analysis of the spectra of HD 192163 and HD 191765. Next Castor and van Blerkom consider the relative intensities of high members of the Pickering series, and from the fact that the populations of the upper levels of high quantum number (i.e., those with $n \geqslant 12$) appear to be in LTE ratios with respect to the number of He^{+2} ions at $T_e \approx 10^5$ K, they deduce that the electron density must be of the order of 5.0×10^{11} cm^{-3} if helium is the dominant constituent of the mantle. A similar density is obtained if one allows for the presence of electrons from hydrogen. A crude check on the significance of the results is obtained by estimating the optical depth in electron scattering.

By putting $\tau_e = N_e \sigma_e R$, it is found that $\tau_e = 1.5$. Castor and van Blerkom conclude that this value is appropriate, but we have seen above that it is too large. The optical depth in electron scattering in the mantle should be less than or of the order of 0.25. It seems that the estimates of N_e and of R may be too large. Another check which Castor and van Blerkom perform is to estimate the intensities of some lines of the He II $(5 - n)$ series. In all cases, their estimated equivalent widths relative to the equivalent width of He II 4686 are too large, which is another indication that the parameter R may be too large.

Castor and van Blerkom complete their coarse analysis of the spectrum of HD 192163 by interpreting the apparent strength of the H I and He II blend at 4100 Å in terms of the relative abundance of hydrogen to helium. They assume that the populations of the upper levels of both He and He II 4100 have LTE ratios with respect to the numbers of H^+ and He^{+2} ions, and they neglect the presence of an underlying hydrogen He line in the photospheric spectrum. We have noted above that proceeding in this manner will lead to an underestimate of the abundance of hydrogen. The calculations of Bhatia and Underhill (1986) show that the departures from LTE population ratios may be large for both H and He^+.

Finally, Castor and van Blerkom (1970) solve the equations of statistical equilibrium using a model He^+ ion consisting of 30 bound levels and a continuum, and adopting the values, $T_c = 4 \times 10^4$ K, $R_c = 13\, R_\odot$, $R_e = 40\, R_\odot$, and $v(R_e) = 1000$ km s^{-1}. They make model mantles corresponding to several combinations of T_e and helium density, N(He). The possible presence of electrons from hydrogen is ignored. Radiative transfer in the model mantle in continuum frequencies is taken into consideration. Castor and van Blerkom show that plausible He II line intensities are obtained if $T_e = 10^5$ K and $N(\mathrm{He}) \approx 2.5 \times 10^{11}$ cm^{-3}. They are able to rule out helium densities less than 10^{11} cm^{-3}, and $T_e > 2 \times 10^5$ K. Their final conclusion is that a radially expanding mantle around a photosphere provides a good model for a Wolf-Rayet star. The treatment is incomplete in the sense that no discussion is given of how the high mantle electron temperature arises or of what generates the outflow velocity. Unlike Rublev (1964b), Castor and van Blerkom make no attempt to account for the energy needed to provide the large number of He^{+2} ions which are implied to be present in the mantle. They just postulate that

He^+ ions are present at the required velocity of outflow and that the energies of the electrons are sufficient to produce the required non-LTE and LTE level population ratios to account for the full observed He II spectrum of WN6 stars. The theory of radiative transfer used in this work is fully described in Mihalas (1978), Chapter 14.

The method of Castor and van Blerkom (1970) uses the radiative-transfer theory developed by Castor (1970) for an atmosphere containing a velocity gradient. In Castor's theory, it is assumed that all sources of continuous opacity except the opacity in continua of the model atom are negligible. We have seen that N_e and the typical path length adopted by Castor and van Blerkom result in an opacity in electron scattering which violates this basic assumption of Castor's radiative-transfer theory.

The early applications of the method of Castor and van Blerkom are listed in Table 7-8. More recently Bhatia and Underhill (1986), using solar composition, have applied the radiative-transfer theory of Castor (1970) in the way outlined by Castor and van Blerkom (1970) with electron densities equal to or less than 10^{10} cm^{-3}. Such electron densities result in negligible optical depth in electron scattering. The relative energies in lines of H, He I, and He II predicted by Bhatia and Underhill agree with the line strengths observed in the case of four Wolf-Rayet stars.

5. Discussion. Rublev's treatment of line formation in Wolf-Rayet atmospheres must be rejected because it postulates a temperature which decreases outward through the mantle, starting from a very hot photosphere. Studies of radiative transfer in moving atmospheres have shown that such a temperature law will not result in emission lines having profiles like those observed for Wolf-Rayet stars. In addition, the observed shapes of the continuous spectra of Wolf-Rayet stars correspond to effective temperatures of the order of 3×10^4 K, rather than 10^5 K. Spectrophotometric observations made from 1250 to at least 8000 Å exist for many Wolf-Rayet stars (see Underhill, 1983b), and none of these observations indicate that the effective temperatures of Wolf-Rayet stars are exceptionally high, near 10^5 K. Rublev's results are valuable, however, because they give the only existing quantitative estimates of how much energy is needed to produce spectra typical of Wolf-Rayet stars. That energy must come from nonradiative sources.

The studies which use the method of Castor and van Blerkom (1970) indicate rather consistently that the electron density in Wolf-Rayet mantles may be typically of the order of a few times 10^{11} cm^{-3} (see Table 7-8, which summarizes the results of these studies). The estimated optical depth due to electron scattering in the mantle is unsatisfactorily large in most cases. This may be because the parameter, R, and hence the parameter, R_e, is systematically overestimated. No check has been applied in any case to determine if the adopted value for R_e is truly appropriate.

The most uncertain factors going into the estimate of the full extent of the line-forming region for He II, the parameter R, when it is estimated in the way proposed by Castor and van Blerkom, are the observed equivalent widths of the He II lines at 3203, 4686, and 10124 Å. The equivalent width assigned to the line at 10124 Å is a particularly sensitive factor. It is straightforward to show that, if the adopted value for the equivalent width for He II 10124 has been overestimated by 30 or 40 percent, the resulting value for R may have been overestimated by about the same amount. Alternatively, R may be estimated from the apparent size of the line-forming region in a well-studied eclipsing binary which contains a Wolf-Rayet star as one component. The range in results for the same star shown in Table 7-8 demonstrates the effect of changing R_e on the properties of an acceptable model mantle. The value of R_e/R_c determines the size of the dilution factor, and this factor is significant for determining the number of photons which are lost through hitting the core of the star, as well as for transmitting the continuum radiation through the mantle. Consequently, changing R_e/R_c will affect the solution of the equations

Table 7-8
Results of Analyses of Wolf-Rayet Spectra

Authors	R_e/R_c	R_c ($R_\odot$)	N_e (cm^{-3})	τ_e*	Stars Studied	Spectral Type
Castor and van Blerkom (1970)	3.1	13	5×10^{11}	0.89	191765	WN6
	3.1	13	5×10^{11}	0.89	192163	WN6
van Blerkom and Patton (1972)	5.8	6	1×10^{11}	0.16	50896	WN5
Castor and Nussbaumer (1972)	3.6	15	4×10^{11}	0.98	γ^2 Vel	WC8
Oegerle and van Blerkom (1976a)	3.0	30	1×10^{11}	0.13	MR119	WN8
Willis and Wilson (1978)	3.75	8	4×10^{11}	0.53	50896	WN5
	3.75	8	4×10^{11}	0.53	191765	WN6
	3.75	8	4×10^{11}	0.53	192163	WN6
Willis and Wilson (1978)	4.0	15	2×10^{11}	0.53	192103	WC8

*Calculated for a path length equal to R_e.

of statistical equilibrium and hence the determination of the source function for each line.

Willis and Wilson (1978) have deduced relative abundances of C/He and N/He in Wolf-Rayet atmospheres. They find that N/He ≈ 0.024 in three WN stars and C/He ≈ 0.009 in one WC star. These values are quite uncertain because Willis and Wilson use schematic model atoms in which they ignore entirely the photoionization and recombination of C and N ions. Similar analyses for additional stars have been made by Smith and Willis (1982, 1983) using the same method. They obtain similar results.

The compositions found by Willis and Wilson, Smith and Willis, and Nugis (1975) are not similar to the composition of cosmic material. The cosmic abundance ratios by number (Allen 1973) are H/He = 11.7, C/He = 0.0039, and N/He = 0.0011. All of the analyses listed in Table 7-8 start with the statement that the abundance of hydrogen is small or negligible. This premise is not securely based (see Bhatia and Underhill (1986)), where it is shown that the NLTE behavior of H may not be ignored, as is done in the above studies. In addition, it is important to consider levels of He I as well as levels of He II in the model He atom.

We must conclude that the information about the composition of Wolf-Rayet atmospheres resulting from the studies listed in Table 7-8 is extremely uncertain. Not only has the analysis for the presence of H been done in a very simple manner, but the analysis of all other ions is severely limited by the very simple model atoms used to determine the coefficients in the equations of statistical equilibrium. Since we know that transitions involving doubly excited states are significant for causing some lines to go into emission in Of stars, it seems inadvisable to ignore the presence of doubly excited states when defining the model atoms for He, C, and N. It is also inadvisable to ignore photoionization from excited states. Only in the analysis of Bhatia and Underhill (1986) is use made of detailed, *realistic* model atoms.

Bhatia and Underhill (1986 and ongoing work) have found that the observed relative intensities of lines of H, He I, He II, C II, C III, C IV, N III, N IV, and N V in six Wolf-Rayet stars can be predicted successfully using cosmic abundances and reliable atomic physics. There is no reason to conclude that the composition of the mantle of any Wolf-Rayet star is anomalous or different from that of an O or Of star. The unusual intensity patterns of the emission lines in Wolf-Rayet spectra appear to be the result of an unusual physical state in the stellar mantle. Bhatia and Underhill find that in WC mantles $5 \times 10^4 < T_e < 10^5$ K and $N_e \approx 10^{10}$ cm^{-3}, while in WN mantles $10^5 < T_e <$

2×10^5 K and $10^9 < N_e < 10^{10}$ cm^{-3}. They confirm that the effective temperatures of Wolf-Rayet stars are in the range found by Underhill (1983b), namely 2.5×10^4 to 3.0×10^4 K. A review of the theories used to interpret Wolf-Rayet spectra can be found in Underhill (1986).

D. The Physics of Winds

In the section on the *Properties of Spherical Flow*, we reviewed the results of the theory which has been used to account for the winds from hot stars. None of the published theories of the winds of hot stars takes into account the possibility of a simultaneous deposition of nonradiative energy and momentum in the region around the critical point. Such deposition seems necessary if the gas is to be accelerated through the critical point to high velocities. Much gas moving at high velocities is observed for O and Wolf-Rayet stars. In this section, we shall make some brief remarks in order to indicate in what directions further study might profitably be done and to place in context the studies of winds driven by the momentum absorbed in lines which have been presented by Castor et al. (1975), Abbott (1978, 1980, 1982), and Weber (1981).

1. Effects of Magnetic Fields. In the mantles of hot stars, as in the mantles of cool stars, energy can be transferred to the gas as a result of the passage of radiation and of the dissipation of waves, acoustic or MHD. Momentum can be transferred to the gas: (1) from the radiation field (see the section on the *Mechanical Force Exerted by Radiation*), (2) as a result of pressure gradients which may exist from the action of global forces like the attraction of gravity reduced by the centrifugal force arising from the rotation of the star and the force due to radiation, (3) as a result of pressure gradients caused by the presence of waves, and (4) by the pressure due to the presence of parcels of hot gas.

That the driving force for the solar wind was the pressure from a layer of very hot gas was suggested by Parker (1958). The action of this force in the atmosphere of a hot star has been explored by Hearn and Vardavas (1981a, 1981b) and by Vardavas and Hearn (1981). The effects of radiation pressure in lines alone as a driving force are described in the next section. The effects of the presence of a magnetic field on the winds from rotating hot stars have been considered by Nerney 1980), Barker (1982), and Barker and Marlborough (1982). Barker and Marlborough show that, when weak magnetic fields are present and there is radial flow, the circular velocity due to rotation may be enhanced or diminished according to the circumstances.

The heating and propulsion of an atmosphere by waves is a subject that has not been explored for hot stars. Let us present a few words on this subject before reviewing the theoretical studies of winds driven by the force of radiation acting in lines. An important difference between the conditions in the mantles of hot and cool stars is the relative importance of wave-related forces and heating and radiation-related forces and heating. Since we have seen that the action of radiation and gravity alone are not sufficient to account for the properties of the winds of O and Wolf-Rayet stars, it is time to consider what the presence of waves may do.

Because convection is not a significant means of transporting energy through the subphotospheric layers of O and Wolf-Rayet stars in comparison to radiation, it is probable that simple acoustic waves generated by turbulence (perhaps resulting from differential rotation of the star), may not be important as a heating agent. In fact, acoustic waves do not seem to be the chief source of energy even for heating the solar corona (see Stein, 1981; and Wentzel, 1981). It must not be overlooked that small magnetic fields may be present in the subphotospheric layers of O and Wolf-Rayet stars (Underhill, 1983a; Underhill, 1984a; Underhill and Fahey, 1984). Then Alfven waves will be generated if turbulence is present, and magneto-acoustic waves may propagate in slow or fast modes. The topic of the generation of propagating waves in stellar atmospheres by turbulence has been discussed by Stein (1981).

Magnetic fields are very important for controlling the state of the mantle of the Sun (see almost every chapter of Jordan, 1981). It is certain that they will have some effect in the mantles of O and Wolf-Rayet stars. The chief question is how important is this effect. Observations of the Sun have shown that magnetic fields can be important in four areas:

1. They may control the generation of the nonthermal energy flux which provides an essential heat input into the mantle.

2. They may control the flow of nonthermal energy by channeling and focusing the propagation of waves.

3. They may control the dissipation of nonthermal energy by allowing plasma instabilities to occur, which affects the dissipation of waves.

4. They may control the rate at which matter and kinetic and internal energy is lost from a star. This is because mass loss occurs preferentially along open magnetic field lines.

Stein has developed scaling formulas by means of which to estimate the flux in acoustic waves in the case no magnetic fields are present, and of Alfvén waves, slow acoustic waves, and magnetic fast acoustic waves in the cases of weak or equipartition turbulent magnetic fields and of strong turbulent magnetic fields. His scaling formulas are valid for the case that, at the base of the photosphere, the opacity coefficient can be written as

$$\varkappa = \varkappa_0 P^a T^b. \tag{7-66}$$

Stein discusses the meaning of the results for the case that $a = 0.7$, $b = 5$, which is appropriate when H^- is the chief source of opacity, and he shows that his scaling formulas probably are valid for late-type stars.

In the case of O and Wolf-Rayet stars, electron scattering is the dominant source of opacity in the atmosphere. In that case, one should use Equation (7-66) with $a = 1$, $b = -1$. Then the ratio of the flux of simple acoustic waves to the total stellar flux varies as $g^{-5/6}T^9{}_{\text{eff}}$, and the relative fluxes carried in the different types of waves vary as follows with respect to the total stellar flux:

Type of Wave	Weak Turbulent Magnetic Field	Strong Turbulent Magnetic Field
Alfvén	$g^{1/12}T_{\text{eff}}{}^{63/12}B^{-1}$	$g^{1/4}T_{\text{eff}}{}^{31/4}B^{-3}$
Acoustic slow	$g^{-1/6}T^5{}_{\text{eff}}$	$T_{\text{eff}}{}^{15/2}B^{-3}$
Acoustic fast	$g^{5/12}T_{\text{eff}}{}^{123/12}B^{-5}$	$g^{5/4}T_{\text{eff}}{}^{91/4}B^{-15}$

Here B measures the typical magnetic field.

The dependence of these expressions on g is about the same as for the ratios given by Stein for the case of opacity due to H^-. However, the dependence on effective temperature is much more steep, except for the case of simple acoustic waves and for acoustic slow waves in the presence of strong turbulent magnetic fields. According to the above scaling laws, the amount of energy to be found in fast acoustic modes, for a given value of g and B, increases much more rapidly with T_{eff} than is the case for Alfvén waves or for acoustic slow modes. Among the O and Wolf-Rayet stars, the range in g is not well known, but it does not appear to be greater than by a factor 2 or 3, which is quite different than in the case of cool stars. The effective temperature may change by at most a factor 3 when going from the coolest objects to the hottest ones (see Underhill, 1982, 1983b). Because of the dependence on T_{eff}, it seems that the amount of energy to be expected in waves in the mantles of O and Wolf-Rayet stars may be far from negligible. To obtain the actual amount, one would have to evaluate the constant factors in the scaling formulas and make an estimate of the average value for the magnetic field.

The amount of nonthermal wave energy flux produced by turbulence depends on the turbulent velocity amplitude, the density at the top of the convection zone, the magnetic field strength, and the fraction of the stellar surface covered by magnetic flux tubes. The O and Wolf-Rayet stars may differ greatly among themselves with respect to the last parameter. Consequently, one may expect quite disparate mantles for stars of rather similar mass, effective temperature, and radius (see Underhill, 1984b). Conversely, one may say that similar mantles need not correspond to similar photospheres. The fraction of the stellar surface covered by magnetic flux tubes appears to be a significant parameter for determining the state of the mantle, thus of determining the major properties of the spectrum which is seen (Underhill and Fahey, 1984).

We have indicated that it is plausible to expect turbulence to be present in the atmospheric layers of O and Wolf-Rayet stars. That substantial "supersonic turbulence" is present in the atmosphere of the sharp-lined B0 V star, τ Sco, has been suggested by Lamers and Rogerson (1978), and its presence is confirmed by Hamann (1981). Hamann (1980) has estimated a "microturbulence" of 100 km s^{-1} for ζ Pup. Our analysis of the line profiles of the sharp-lined O star, 10 Lac (see the section on *Analyses of the Absorption Lines of O Stars: Abundances*), has similarly indicated the presence of a velocity field which broadens intrinsically strong lines from abundant atoms symmetrically by a considerable amount. Therefore, it is certain that a large field of randomly directed velocities exists in the layers in which the strong absorption lines of O stars are formed. This velocity field, acting on gas in the presence of magnetic fields, could generate Alfvén waves and acoustic waves in the slow and the fast modes as postulated in the discussion by Stein (1981). There are no observations giving direct information on the presence or absence of similar randomly directed fields of motion in the atmospheres of Wolf-Rayet stars, but the fact that most of the emission lines have Gaussian shapes is evidence in favor of such velocity fields being present.

2. Models for Radiatively Driven Winds Powered by Absorption in Lines. The flow of a wind from a star must take place according to the equation of motion which may be derived from the conservation equations (see the section on *Conservation Laws*). Castor et al. (1975) worked out the consequences of introducing into Equation (7-17) a force due to radiation, having the form of Equation (7-32), including the approximation expressed in Equation (7-33), to act against the force due to gravity. They specify that the temperature law in the atmosphere, $T(r)$, is that for an LTE model atmosphere in radiative equilibrium (Castor, 1974a) over the range of radii in which hydrostatic equilibrium is a good approximation. Between the photosphere and the sonic point, the temperature is postulated to increase from 38000 to 50000 K, but no energy input is specified to cause this heating. Near the singular point, which is found to lie at about 1.5 to 2 times the photospheric radius, it is assumed that the temperature decreases according to a power law in r. It is argued that the exponent of this power law lies between 0 and 0.5; further, it is stated that the sonic radius is essentially the same as the photospheric radius, since the density rises very rapidly interior to the sonic point. Under these conditions, the equation of motion for the wind is solved, and a density pattern is found using the assumption of spherical flow expressed in Equation (7-19).

The weakness of the treatment by Castor et al. (1975) is that the expression which they adopt for the force due to radiation is valid only when large velocity gradients are present. They use this expression throughout the model atmosphere, including in the regions in which the velocity of flow is small. Castor et al. give no justification for the assumption that the velocity gradients needed to validate their use of Equation (7-32) exist. Abbott (1978, 1980, 1982) has expanded the treatment of Castor et al. by working out analytical expressions for the force due to the absorption of radiation in lines. He continues to use Equation (7-32), assuming implicitly that the required velocity gradients needed for its validity are present. In the second paper, Abbott presents a physical interpretation of the wind problem solved by Castor et al. Here Abbott shows that the presence of rotation and of coronal gas will change the velocity law in real stars somewhat from the result attained by Castor et al. In his papers, Abbott discusses how many lines need to be considered and what the intrinsic strengths of these lines may be. It should be noted that only resonance lines and a few strong lines arising from low-lying metastable levels are observed to show wind profiles in the spectra of O stars. This is in contrast to Abbott's claim that many subordinate lines contribute to the accelerating force caused by the absorption of radiation.

Weber (1981) has examined the question of how winds may be driven by radiation pressure in lines when the flow velocities are of the order of the sonic velocity. He shows that the approximation for the force due to radiation used by Castor et al. (1975) overestimates the acceleration near the sonic point. Weber goes on to study the hydrodynamic equations for the wind in dimensionless form, and he obtains solutions for several values of the parameters which he defines to describe the situation in the vicinity of the critical point and elsewhere in the wind. Although Weber is able to find solutions which give stellar spectra like those which are observed to exist for O stars, a study of the parameters he needs in order to obtain a wind solution shows that the key parameter which describes the opacity in the lines must take values which, at the critical point, are unlike the values expected from slowly moving material of cosmic composition at the densities and temperatures which occur in the atmospheres of O stars. In Weber's model mantles, the wind temperature is assumed to be that appropriate for radiative equilibrium in a stationary spherical gray atmosphere.

Weber's systematic exploration of the properties of winds driven by radiation absorbed in lines is useful, because it confirms that unrealistic values for the line-absorption coefficients must be assumed if a solution of the wind equation is to be obtained without adding momentum and energy to the gas from sources other than the radiation field.

In the studies by Castor et al. (1975), Abbott (1978, 1980, 1982), and Weber (1981), it is assumed that the gas in the mantle is illuminated by a *continuous* spectrum. This cannot be the case, particularly at deep layers in the mantle near the sonic point, because the photospheres of O stars produce absorption lines. Many lines from the third and fourth spectra of C, N, O, Si, and the metals are seen in the ultraviolet and visible spectra of O stars; in the hottest stars, lines are also seen from the fifth and sixth spectra of N, O, and the metals. Most of these lines come from excited levels, and they are not displaced by a large amount. That the resonance lines of these spectra are formed in significant strength in the photosphere is a foregone conclusion if subordinate lines are seen at moderate strength. That photospheric C IV resonance lines of considerable strength are formed in the photospheres of middle and late luminous O-type stars can be deduced from the fact that the wind profiles of the C IV lines have very weak or absent emission components (see Underhill, 1983a). In Underhill and Doazan (1982), Chapter 4, it was pointed out that the absence or great weakness of emission components for the wind lines of Si IV in the spectra of middle B supergiants could be understood as being due to the presence of *photospheric* Si IV absorption lines. Castor and Lamers (1979) have done calculations using an ad hoc model atmosphere to illustrate that, if a strong photospheric absorption line is present, the wind profile formed in the mantle of the same line will lack strong emission. Because the photospheric spectrum which is supposed to be driving the wind in the theories of radiatively driven winds contains absorption lines, the true value of the force of radiation will be even less than it is assumed to be in these theories.

The formulas which are used in the theories of winds driven by absorption in lines (Equations (7-32) and (7-33)) to find the force due to radiation are valid only when a sufficient speed of outflow has been established (by other means) to displace the absorption lines formed in the mantle shortward of significant photospheric absorption lines and to set up significant velocity gradients. Once this has occurred, radiation can become a strong driving force. The high outflow velocities shown by early-type stars are, at least partially, the result of driving by radiation pressure; however, the force of radiation is not the factor which initiates the wind.

The winds of the early-type stars may be like the winds of the late-type stars in requiring magnetohydrodynamic causes for their origin. They differ from the winds of late-type stars in that they may be accelerated to high velocities by means of the force of radiation. The terminal velocities of the winds from early-type stars are related to the total amount of radiation available to accelerate the gas in the mantle; there is a loose correlation between T_{eff} and v_{∞} (Underhill, 1983a, 1983c). On the other hand, if winds are initiated by magnetohydrodynamic events, the rate of outflow of matter per square centimeter of surface on the star will be determined by the local density of magnetic flux tubes open to interstellar space. This latter quantity may vary from time to time in the case of any one star. Indeed, the observations of variable discrete components in wind resonance lines indicate that it does do so. Stars of the same spectral type may then be expected to have similar winds only so far as they have similar distributions of surface magnetic field (Underhill and Fahey, 1984).

V. RATE OF MASS LOSS FROM HOT STARS

All published estimates of the rate of mass loss from hot stars are based on the assumption that the flow is spherically symmetric. The rate of mass loss is found by means of Equation (7-19). If propulsion by magnetic fields initiates the outflow, however, uniform outflow is unlikely to occur.

A. Methods for Finding the Rate of Mass Loss

Five different methods have been used for estimating $\varrho(r)$ and $v(r)$ at a particular radius r and for finding $\dot{M}$ from Equation (7-19). They are as follows:

1. By combining an observed velocity of outflow and the radius of the star with the density in the stellar atmosphere, estimated from an analysis of parts of the spectrum.

2. By interpreting the radio flux from the star in terms of simple theory.

3. By interpreting the infrared excess of the star in terms of simple theory.

4. By interpreting the shapes and strengths of resonance lines formed in a wind by means of one of the applications of the theory of line formation in an expanding atmosphere.

5. By accounting for the equivalent width and/or the profile of Hα in terms of one of the applications of the theory of line formation in an expanding atmosphere.

Method 1 was used to obtain the first estimates of the rate of mass loss from Wolf-Rayet stars (see, for instance, Kosirev, 1934; Underhill, 1968b). It is clear now that these first estimates give an upper limit to the rate of mass loss from Wolf-Rayet stars because a density and a radius, each of which is suitable for the photosphere, are combined with a velocity of outflow equivalent to that which can occur only high in the mantle where the density is less than what it is in the photosphere. If one combines a photospheric density and radius with an estimate of the outflow velocity at the sonic point, one may obtain a more accurate estimate of $\dot{M}$. Typically, in the case of O and Wolf-Rayet stars, the sonic velocity may be of the order of 2 percent of the outflow velocities which have been used in the applications of Method 1.

Methods 2 and 3 have been described in the section on the *Radio and Infrared Flux from Spherical Atmospheres: Mass Loss.* A maximum outflow velocity determined from the observed absorption troughs of resonance lines is combined with the density and radius estimated from considering the amount of gas required to provide the observed amount of free-free emission at either a radio or an infrared wavelength. It is assumed that the gas fills a large spherical volume around the star as a result of uniform spherical outflow. However, uniform spherical outflow is not the only way to maintain a body of gas around a star. Some of the gas might be suspended around the star in large magnetic loops. Suspension in magnetic loops is what occurs for the Sun, and it permits a large radio flux to be radiated when the rate of mass loss is very small.

The fact that the radio spectrum and much of the infrared spectrum from luminous early-type stars fit a power law in wavelength indicates that a density gradient is present in the optically thin parts of the emitting gas from which the excess radiation comes. The observed power-law character of the excess spectrum does not prove that the density gradient is due to outflow at a constant velocity. This point has been noted by Seaquist (1976).Some theoretical arguments about what may happen when low-density material is suspended near a hot star in magnetic loops has been given by Underhill (1983a, 1984a).

Method 4, the interpretation of ultraviolet resonance line profiles, was applied first to the spectrum of ζ Pup by Lamers and Morton (1976). Since then, Method 4 has been applied to the spectra of many O stars by several groups of observers. The results are summarized in the next section. Castor et al. (1982) have devised a formulation of the theory which may be used to interpret the profiles and equivalent widths of lines formed in the mantle when they are measured on low-resolution ultraviolet spectra. They have applied this theory to finding the rate of mass loss from the central star of the planetary nebula NGC 6543. All the other applications of Method 4 make use of high-resolution ultraviolet spectra.

Method 5, interpretation of the emission feature at Hα was applied first by Hearn (1975a) and then by Cassinelli et al. (1978) to find a rate of mass loss from the equivalent width and profile of the Hα emission in the spectrum of the O9.5 Ib star, ζ Ori A. A version of the method based on the theory of radiatively driven winds was developed and applied by Klein and Castor (1978); this formulation has been applied also by Conti and Frost (1977). The problem of interpreting the profile of Hα in the spectrum of an O star or a B-type supergiant in terms of $\dot{M}$ has been studied by Olson and Ebbets (1981), who have given results for six O stars.

The rates of mass loss found by these five methods do not always agree closely for any one star. This is because each method makes use of different observational data and each is therefore sensitive to different factors in the rather simple models which are used to represent the physical state of the wind and the process of spectrum formation which takes place in the wind of the star. The differences found between the values of $\dot{M}$ obtained in the different ways are potentially interesting because they may shed light on the validity of the models which have been used.

B. Estimated Rates of Mass Loss from O and Wolf-Rayet Stars

The rates of mass loss from O stars which have been estimated from various observed quantities are listed in Table 7-9. The stars are in order of right ascension. The individual papers should be consulted to determine the steps followed in each method and to determine what assumptions have been made concerning the temperature in the mantle, the velocity law, and the state of ionization of the gas in the wind. In the studies listed in Table 7-9, the models used have been specified in sufficient

Table 7-9
Rate of Mass Loss from O Stars

Star HD/HDE/CPD/BD	Spectral Type	Method	$\dot{M}$ ($10^{-6}\ M_{\odot}\,yr^{-1}$)	Ref.*
108	O7 If	IR	2.4	1
1337 AO Cas	O9 III+O9 III	UV	5.0	2
14947	O5.5f	Hα	8.0	3
		IR	2.4	1
		Hα	9.1	4
		Radio	<24.0	6
		UV	4.0	5
15558	O5 (f)	Hα	5.9	4
		Radio	<15.0	17
15570	O4f	Hα	18.0	3
		Hα	24.0	4
		Radio	<21.0	6
15629	O5((f))	Hα	2.3	4
		UV	2.5	5
16429	O9.5 I	Hα	2.9	4
16691	O4f	Hα	8.0	3
		Hα	8.1	4
30614 α Cam	O9.5 Ia	Hα	3.5	7
		Hα	2.0	7
		Radio	1.7	8
36486 δ Ori A	O9.5 II	IR	1.2	1
		IR	0.92	1
		Radio	<6.0	14
		UV	0.5	9
36861 λ Ori A	O8 III((f))	IR	0.38	1
		Hα	<0.6	7
		UV	0.2	9
269698	O4f	UV	2.5	5
37041 θ^2 Ori A	O9 V	UV	0.064	10
37043 ι Ori	O8.5 III	Hα	<0.9	7
		UV	0.3	9
269810	O3f	UV	5.0	5
37742 ζ Ori A	O9.5 Ib	Hα	1.8	11
		IR	1.2	1
		IR	1.0	1
		Hα	3.2	4
		Hα	3.0	16
		Radio	2.3	6
		Hα	2.3	7
		Hα	1.2	7
		UV	0.5	9
38666 μ Col	O9 V	UV	0.03	9
39680	O9 Ve	IR	1.9	1

*References are listed at end of table.

Table 7-9 (Continued)

Star HD/HDE/CPD/BD	Spectral Type	Method	$\dot{M}$ ($10^{-6}\ M_\odot\ yr^{-1}$)	Ref.*
39680	O9 Ve	IR	1.8	1
42088	O6.5 V	UV	0.13	5
45314	OBe	IR	0.37	1
46056	O8 V(e)	UV	0.025	5
46149	O8.5 V	UV	0.050	5
46150	O5.5((f))	UV	0.79	5
46223	O4((f))	UV	2.5	5
47839 15 Mon	O8 III((f))	Hα	<0.6	7
		UV	0.15	9
48099	O6.5 V	UV	0.63	5
54662	O7 III	UV	0.20	5
57060 29 CMa	O8.5 If	UV	3.0	12
		Radio	6.8	6
60848 BN Gem	O8 Vpe	IR	2.4	1
66811 ζ Pup	O4ef	UV	7.2	13
		Hα	9.0	3
		IR	3.5	1
		Hα	8.0	16
		Radio	6.3	14
		Radio	3.5	6
		Hα	6.0	7
		Hα	5.0	7
		UV	7.0	9
		UV	3.5	5
93129 A	O3f	Hα	21.0	3
93204	O5 V((f))	UV	2.0	5
93222	O7 III((f))	UV	0.50	5
−59°2600	O6 V((f))	UV	1.0	5
93250	O3 V((f))	UV	1.3	5
−59°2603	O7 V((f))	UV	0.063	5
303308	O3 V(f)	UV	2.5	5
101190	O6 V((f))	UV	1.0	5
101298	O6 V((f))	UV	1.6	5
101413	O8 V	UV	0.13	5
101436	O6.5 V	UV	1.3	5
112244	O8.5 Iab	Radio	< 28.0	14
149757 ζ Oph	O9 V(e)	Radio	< 0.38	6
151804	O8 If	Hα	17.0	4
		Radio	< 34.0	14
		UV	7.9	5
152233	O6(f)	Hα	3.7	4
		UV	4.0	5
152247	O9.5 I	Hα	2.7	4
152408	O8 If	Hα	21.0	4
		Radio	< 29.0	14

*References are listed at end of table.

Table 7-9 Continued

Star HD/HDE/CPD/BD	Spectral Type	Method	$\dot{M}$ ($10^{-6}\ M_\odot\ yr^{-1}$)	Ref.*
152408	O8 If	UV	10.0	5
152424	O9.5 I	Hα	5.2	4
162978	O8.5 IIIf	Radio	$<$ 5.4	8
163758	O6.5 Iaf	UV	6.3	5
164794 9 Sgr	O4((f))	Radio	25.0	6
		UV	5.0	5
		Radio	38.0	17
167659	O7 I((f))	UV	3.2	5
167771	O8 I((f))	Hα	2.2	4
−11°4586	O8 I	Radio	$<$14.0	8
168076	O4((f))	Hα	10.0	3
175754	O8 IIIf	Radio	$<$14.0	8
175876	O6.5 III(f)	Radio	$<$18.0	8
188001 9 Sge	O8 If	UV	6.3	5
190429N	O4f	Hα	11.0	3
		Radio	$<$21.0	6
190864	O7 III((f))	UV	2.0	5
Cyg OB2 No. 5	O6f+O7f	Radio	28.0	17
Cyg OB2 No. 7	O3 If	Radio	$\leqslant$19.0	17
Cyg OB2 No. 8A	O6 Ib(f)	Radio	32.0	17
Cyg OB2 No. 9	O5 If	Radio	118.0	17
Cyg OB2 No. 11	O5 If	Radio	$<$13.0	17
207198	O9 I	Radio	$<$ 3.2	8
209481	O8.5 III+O9.5 V	Radio	$<$ 3.1	8
210839 λCep	O6ef	Radio	4.8	6
		Radio	7.8	6
		UV	4.0	5
		Radio	$<$ 5.0	17
217086	O7 V	UV	0.42	15
225160	O8 Ib(f)	Radio	12.0	8

*References:

1. Barlow and Cohen (1977)
2. McCluskey and Kondo (1981)
3. Conti and Frost (1977)
4. Klein and Castor (1978)
5. Garmany et al. (1981
6. Abbott et al. (1980)
7. Olson and Ebbets (1981)
8. Felli and Panagia (1981)
9. Olson and Castor (1981)
10. Bernacca and Bianchi (1979
11. Hearn (1975a)
12. McCluskey et al. (1975)
13. Lamers and Morton (1976)
14. Morton and Wright (1979)
15. Perinotto and Panagia (1981)
16. Cassinelli et al. (1978)
17. Abbott et al. (1981)

detail that numerical values are found for the rate of mass loss without resort to empirical calibrations of the numerical quantities which result from the observed data. Spectral types differing slightly from those listed in Table 7-9 can be found in some lists.

Additional results due to Hutchings and von Rudloff (1980) and to Gathier et al. (1981) are listed in Table 7-10. These results are not fully independent because they make use of calibrations based on some of the results given in Table 7-9. Hutchings and von Rudloff evaluated certain constant terms in their expression for $\dot{M}$ by requiring that the rate of mass loss from ζ Pup be $6.0 \times 10^{-6} M_{\odot}$ yr^{-1}, while Gathier et al. adopted published rates of mass loss from nine stars to find empirically the value of a multiplicative constant in their formulation of the expression for $\dot{M}$.

The mass-loss rates of Hutchings (1976) for 27 O stars are not listed because these values depend upon a very rough empirical calibration of Hutchings' mass-loss indicators in terms of the actual rate of mass loss. Likewise, the results of Conti and Garmany (1980) are not listed. This is because Garmany et al. (1981) have reworked the observational data of Conti and Garmany using an improved method of analysis, and they give results which supersede those of Conti and Garmany.

The uncertainties of the individual values of $\dot{M}$ are difficult to estimate. The authors suggest that the true values for $\dot{M}$ may lie in the range from about 50 percent of the listed values to three times these values. Many of the results obtained by means of radio observations are upper limits set by the detection limit of the

Table 7-10
Rates of Mass Loss from O Stars Estimated from the Ultraviolet Spectrum by Scaling

Star HD/Name	Spectral Type	$\dot{M}$ ($10^{-6} M_{\odot}$ yr^{-1})	Ref.*
108	O7 If	0.2	1
24912 ξ Per	O7.5 III((f))	1.8	2
30614 α Cam	O9.5 Ia	4.2	2
36486 δ Ori A	O9.5 II	1.3	2
36861 λ Ori A	O8 III((f))	1.0	2
37043 ι Ori	O8.5 III	1.1	2
37742 ζ Ori A	O9.5 Ib	0.3	1
		2.3	2
38666 μ Col	O9 V	0.063	2
47839 15 Mon	O8 III((f))	0.85	2
57061 τ CMa	O9 II	1.9	2
66811 ζ Pup	O4ef	6.0	1
		4.1	2
93129A	O3f	2.0	1
148937	O6.5f	0.2	1
149038 μ Nor	O9.7 Iab	2.0	2
149757 ζ Oph	O9 V(e)	0.32	2
152408	O8 If	13.0	1
210839 λ Cep	O6ef	5.0	1

*References:
1. Hutchings and von Rudloff (1980).
2. Gathier et al. (1981).

radio telescope used. Some results may be biased more strongly by the assumptions underlying the theory used than others are, but it is not clearly evident that this is so. In particular, it is not a priori certain that the results obtained for O stars from the analyses of radio fluxes are more secure than those obtained by other methods which make use of other assumptions about the size of the emitting volume, the outflow velocity, the electron temperature in the wind, and the composition of the wind. All of the methods assume uniform spherical outflow.

Garmany et al. (1981) have estimated that the rates of mass loss found by their use of the profiles of lines formed in a wind computed by Castor and Lamers (1979) are accurate to within a factor of 2 in most cases. This will be true only if the degree of ionization in the wind changes outward in the way that is implied by the dependence on radius of the function used by Castor and Lamers to describe how the ions forming the resonance line are distributed in space. Evaluation of the adopted function with use of the values of the parameter γ which are inferred by Garmany et al. shows that a rapid decrease outward of the number of ions was implied. If, on the other hand, the degree of ionization of the gas in the wind remains constant throughout much of the wind (as is observed to be the case for the Sun and is believed to be due to the deposition of energy in the solar wind by a mechanism which is active far from the surface of the Sun), then the true value of $\dot{M}$ might be significantly lower than the value estimated by Garmany et al. The true values of $\dot{M}$ for O stars may be smaller than the values given in Tables 7-9 and 7-10 by a factor as large as 10. They are unlikely to be larger than these values.

The published rates of mass loss from Wolf-Rayet stars are collected in Table 7-11. The spectral types are from the catalog by van der Hucht et al. (1981). The stars are listed in order of right ascension. All of these results have been obtained by interpreting measured or estimated radio fluxes by means of the theory of Wright and Barlow (1975). Those studies which make use of a radio flux found by projecting a power-law spectrum with slope α from an observed infrared excess at 10 μm are noted by the designation IR in the column indicating the method. Underhill (1980a) assumed that α = 0.60, while Barlow et al. (1981) adopted α = 0.76. Even though the same observational data are used for some stars in these two papers, the authors did not always start from the same value for the infrared excess at 10 μm. They make different estimates of the parameters describing the emitting plasma, the distance of the star, and the constant velocity of outflow. Different choices for all these quantities may modify the derived values of $\dot{M}$ by a factor of 4 to 8. The word *radio* is used in the column indicating the method when observed radio fluxes have been used.

Each published result is presented in Table 7-11 so that the reader may see the range in $\dot{M}$ which results from responsible, but different, estimates of the values of the quantities needed to convert an observed radio flux or infrared excess into a rate of mass loss by means of Equation (7-39). Although it was pointed out by Seaquist (1976) that the fact that the infrared excesses and radio fluxes from Wolf-Rayet stars may be connected by a power law does not necessarily imply that the emitting plasma is solely due to uniform spherical outflow at a constant velocity, this description of the emitting plasma has been adopted universally. The rates of mass loss from individual Wolf-Rayet stars deduced by Hackwell et al. (1974) from their observed infrared excesses are not given in Table 7-11 because these values have been found by primitive theory. An outflow of 1000 km s^{-1} was adopted for each star by Hackwell et al. The rates of mass loss found by Hackwell et al. fall in the range from 0.6 to 2.8 $\times$ 10^{-5} $M_{\odot}$ yr^{-1}. No entry is made in Table 7-11 for HD 193793 because the radio flux and infrared excess of HD 193793 have been observed to vary.

A systematic difference exists between outflow velocities estimated only from lines in the visible part of the spectrum and the outflow velocities estimated by means of the absorption

Table 7-11
Estimated Rates of Mass Loss from Wolf-Rayet Stars

HD	Other Name	Spectral Type	v_∞ (km s^{-1})	$\dot{M}/v_\infty$ [a]	Method	$\dot{M}$ (10^{-6} $M_\odot$ yr^{-1})	Ref.[b]
4004		WN5	2600	9.6	IR	25	1
16523		WC5	2800	<10.7	Radio	<30	2
50896		WN5	1400[c]	27.1	IR	38	3
			2700	12.2	IR	33	1
			2700	11.1	Radio	31	4
			2700	11.9	Radio	32	4
56925		WN4	2400	<25.0	Radio	<60	4
68273	γ^2 Vel	WC8 + O9 I	1000[c]	12.0	Radio	12	5
			2900	13.4	Radio	39	6
			2900	17.9	Radio	52	7
			2000	29.5	IR	59	1
92740		WN7 + abs	2600	14.6	IR	38	1
93131		WN7 + abs	2800	12.1	IR	34	1
93162		WN7 + abs	2900	11.7	IR	34	1
113904	θ Mus	WC6 + O9.5 I	3500	9.7	IR	34	1
151932		WN7	2250	17.3	IR	39	1
			2250	18.2	Radio	41	2
			2250	19.6	Radio	44	4
152270		WC7 + O5–8	3300	14.8	IR	49	1
			3300	4.2	Radio	<14	2
			3300	15.2	Radio	50	4
165688		WN6	2400	12.5	IR	30	1
			2200	10.9	Radio	24	2
165763		WC5	1900[c]	29.5	IR	56	3
			3700	11.6	IR	43	1
			2800	8.6	Radio	24	2
168206	CV Ser	WC8 + O8–9	1800	<10.0	Radio	<18	2
	AS 320	WC9	2200	<12.7	Radio	<28	2
	BAC 209	WN8	1800	11.1	IR	20	1
	AS 374	WN8	1800	28.3	IR	51	1
			1400	< 6.4	Radio	< 9	2
	MR 97	WN7 + abs	2600	< 1.2	Radio	< 3	4
190002		WC6	2600	< 3.5	Radio	< 9	2
190918		WN4.5 + O9.5 Ia	1700	11.2	IR	19	1
			1700	<11.8	Radio	<20	4
191765		WN6	1600[c]	36.2	IR	58	3
			2600	12.7	IR	33	1
			2600	10.4	Radio	27	4
			2600	14.6	Radio	38	4
192103		WC8	1200[c]	30.0	IR	36	3
			2000	13.5	IR	27	1
			2000	11.5	Radio	23	4
192163		WN6	1400[c]	25.7	IR	36	3

Table 7-11 (Continued)

HD	Other Name	Spectral Type	v_∞ (km s^{-1})	$\dot{M}/v_\infty$[a]	Method	$\dot{M}$ ($10^{-6}\,M_\odot$ yr^{-1})	Ref.[b]
192163		WN6	2250	10.2	IR	23	1
192641		WC7 + abs	2600	5.0	Radio	13	2
193077		WC5 + abs	1700	10.0	IR	17	1
			1700	5.9	Radio	10	2
193576	V444 Cyg	WN5 + O6	2500	7.2	IR	18	1
			2500	5.6	Radio	14	2
	MR 110	WC5	2800	10.0	Radio	28	2
	MR 111	WN-WC	2600	5.8	IR	15	1
	MR 112	WC4	3700	8.6	IR	32	1

[a]Units are $10^{-9}\,M_\odot$ yr^{-1} (km s^{-1})$^{-1}$.

[b]References: 1. Barlow et al. (1981) 4. Hogg (1982) 6. Morton and Wright (1978)
2. Abbott et al. (1982) 5. Seaquist (1976) 7. Morton and Wright (1979)
3. Underhill (1980)

[c]Outflow velocity estimated from visible spectra.

troughs seen in the ultraviolet part of the spectrum. The values of v_∞ found by means of the ultraviolet spectrum are probably more representative than the values from the visible spectrum used by Seaquist (1976) and Underhill (1980a).

The values of $\dot{M}/v_\infty$ which are deduced for Wolf-Rayet stars depend on the assumptions made concerning the physical state and composition of the emitting plasma and the distance of the star. There is no systematic correlation between $\dot{M}/v_\infty$ or $\dot{M}$ and spectral type. If one ignores results which are upper limits, one finds that the average value of $\dot{M}/v_\infty$ with its standard deviation about the mean value is $(14.5 \pm 7.4) \times 10^{-9}\,M_\odot$ yr^{-1} (km s^{-1})$^{-1}$. The average value of $\dot{M}$ is $(32.5 \pm 12.7) \times 10^{-6}\,M_\odot$ yr^{-1}. The standard deviation of $\dot{M}/v_\infty$ corresponds to an uncertainty of 50 percent in $\dot{M}/v_\infty$. This is twice the uncertainty estimated in the section on the *Radio and Infrared Flux from Spherical Atmospheres: Mass Loss.*

The significance of the estimated values of $\dot{M}$ for models of the mantles of O and Wolf-Rayet stars is discussed in Chapter 8. It is pertinent to note here that the large rates of mass loss given in Table 7-11 imply that the *photospheres* of Wolf-Rayet stars are expanding at supersonic rates (see Underhill, 1983a). It is difficult, if not impossible, to justify this circumstance in terms of the physical picture commonly used for representing what is occurring in the atmospheres of hot stars.

VI. ATMOSPHERIC STABILITY: X RAYS

In this section, we shall review studies which have as their goal providing an understanding of and an explanation for observations which cannot be explained by the theoretical concepts which have been reported above. We have seen that the theory of radiative transfer in stationary layers of hot gas allows us to understand and interpret the continuous spectrum from O and Wolf-Rayet stars. In addition, this theory provides a self-consistent interpretation for many of the observed details of the line spectrum of an O star. However, we have seen (in the section *Analyses of the Absorption Lines of O Stars: Abundances*) that in the spectrum of the main-sequence O star, 10 Lac, the leading members of series of lines from abundant atoms and ions are too strong and too broad to be represented by the theory of spectrum formation in hot stationary gas. In order to account for the observed line broadening qualitatively, one may postulate the presence of large randomly directed motions—of "turbulence".

The velocity field which is envisaged is usually called "microturbulence" even though the motions range up to 100 km s^{-1} or so and are clearly supersonic. Lamers and Rogerson (1978) have noted that, to understand the origin of the longward wings of the profiles of the resonance lines of N V and O VI in the spectrum of τ Sco, B0 V, one must postulate that "turbulent" velocities of 150 km s^{-1} are present at the interface between the mantle and the photosphere. Similarly, Hamann (1980) has found it necessary to postulate the presence of "turbulence" of the order of 100 km s^{-1} in order to account for the shapes of the ultraviolet resonance lines formed in the wind of ζ Pup, O4 If. The presence of "turbulence" in the mantles of these stars implies that the flow of the wind is not laminar in the parts of the mantle in which the extreme wings of the spectral lines under discussion are formed. Presumably, this place is close to the photosphere, but this is not necessarily so, owing to the breakdown of the Eddington-Barbier relation in moving atmospheres.

The question whether instabilities which manifest themselves as turbulent motion can be generated by the radiation in the atmospheres of stars has been examined by Spiegel (1976). He develops a mathematical description of the problem and points out the resemblance of the astrophysical problem to the geophysical problem of fluidized beds.

The importance of nonlinear effects for determining the character of the instabilities and resulting velocity fields which may appear in stellar atmospheres has been reviewed by Leibacher and Stein (1976). They point out that the dissipation of shocks arising from instabilities deep in the atmosphere may result in the transfer of energy from the subphotosphere to the mantle by nonradiative means.

The theory of spectrum formation in a wind which flows in a spherically symmetric, laminar manner from a star, and which is in radiative equilibrium, gives an acceptable explanation for many of the details of the absorption and emission lines which are formed in the mantles of O and Wolf-Rayet stars (see the section on *The Line Spectrum from Moving Three-Dimensional Model Atmospheres*). However, laminar flow explains neither the presence of discrete, rather sharp, displaced absorption components of the ultraviolet resonance lines of O stars, nor changes of these components in times of the order of hours to months, nor the facts that: (1) ions requiring much energy for their formation, such as O^{+5}, are seen in the atmospheres of O stars, and (2) X rays are emitted by O stars in small amounts. Underhill and Fahey (1984) have shown that the discrete components may be formed in parcels of gas released from points above the photosphere of the rotating star, while Waldron (1984) has modeled the ionizing action of X rays generated in a thin corona near the photosphere of the star and shown that he can account for the observations.

The observations show that the mantles of O stars are inhomogeneous, and they suggest that the mantles contain parcels of plasma having electron temperatures in the range from 10^6 to 10^7 K. The fact that X rays escape from the mantles of O stars in sufficient quantity to be observed at the Earth suggests that parcels of very hot plasma exist in the outer parts of the winds emanating from O stars.

It was suggested by Mihalas (1969) that the winds from hot stars may be heated and propelled as a result of the deposition of nonradiative energy (released by Rayleigh-Taylor instabilities) in the photospheres of hot stars. Lucy and White (1980) have shown that sufficient nonradiative heating to account for the X ray fluxes which are observed from O stars may be provided in line-driven winds by chaotic hypersonic motions. A critical point for the theory of Lucy and White is whether or not instabilities dissipating mechanical energy are generated in sufficient quantity to provide the needed field of chaotic hypersonic motions.

Lucy (1982) has studied the heating which may be caused by the passage of shocks through the wind.

Clearly, it is necessary to understand how stable the outer layers of a hot star are against disruption by small perturbations in the velocity, temperature, and density of the flow and to know what types of instability may occur. In this section, we shall report research which addresses these questions. It was recognized long ago (Underhill, 1949) that a combination of high temperature, high degree of ionization, large amounts of radiation, and low particle densities would make the atmospheres of the hottest stars susceptible to convection and similar mechanical instabilities. Only recently, however, have attempts been made to describe theoretically what may occur. If magnetic fields are present, plasma instabilities in addition to the Rayleigh-Taylor instability may occur. That instability describes what happens when two liquids of different densities are superimposed, the one on the other, or when the two liquids are accelerated toward each other. Experiments in the laboratory indicate that hydromagnetic instabilities would tend to break the laminar flow into blobs and to heat portions of the gas to high temperatures in the same manner in which Rayleigh-Taylor instabilities are believed to act. However, motion would be constrained in some directions by the presence of a magnetic field. Chandrasekhar (1961), Chapter IV, has shown that all steady slow motions in the presence of a uniform magnetic field are necessarily two-dimensional; motions cannot vary in the direction of the magnetic field.

It is useful to think of the flow from a star as being composed of three parts: (1) a subsonic section in which the flow velocity is significantly less than the ambient velocity of sound in the plasma; (2) a trans-sonic region in which the ratio of the flow velocity to the ambient sound velocity is $1 \pm \epsilon$, where ϵ is a small quantity; and (3) a supersonic region in which the flow velocity greatly exceeds the ambient sound velocity. The geometric extent of the first two regions is probably small relative to the radius of the star, but the geometric extent of the supersonic region may be 50 to 100 times the radius of the star. Most of our spectroscopic observations of winds give information about conditions in the supersonic part of the wind.

We leave aside here any discussion of what may be the physical cause of an initial outflow velocity, be the initial velocity ever so small, and of the cause of the perturbations of the physical state which are assumed to occur. Some statements on these topics can be found at the end of Chapter 8. Any fully satisfactory theory of the mantles of hot stars must provide physical causes for the initial conditions which generate heating and flows in a mantle. None of the published theoretical studies of the stability of the winds from hot stars, and the implications of instability, have yet reached such a stage of completion that they specify causes for the perturbations which are postulated to be present. Likewise, in the case of early-type stars, there has been no discussion of how the ensuing thermal instabilities might be affected by the presence of a magnetic field or by rotation. In Chandrasekhar (1961), Chapters IV and V, the properties of thermal instabilities in the presence of magnetic fields and of rotation are discussed.

A. Stability in the Subsonic and Trans-Sonic Regions

Cannon and Thomas (1977) have studied the stability of a photosphere in the presence of a one-dimensional systematic outward flow. They show that inside the sonic point it is possible to construct a model atmosphere that is in hydrostatic equilibrium but that, in the case of a model for which $T(r)$ and $\rho(r)$ develop in such a way as to match the conditions in the interstellar medium at distances far from the star, any small radial flow is unstable against small perturbations in the flow. Acoustic waves in the flow will amplify into shocks and produce heating as the flow enters the trans-sonic region.

Nelson and Hearn (1978) have noted that a Rayleigh-Taylor-type instability may occur in the trans-sonic region of the wind from a hot star if the driving force for the wind is due to the absorption of line radiation in the presence of a velocity gradient. As a result of this instability, radially symmetric stellar winds driven by resonance-line radiative forces will break up into small horizontal lengths. Nelson and Hearn suggest that the energy fed into the instability will provide a significant source of heating in the trans-sonic region.

Martens (1979) has continued this study of the stability of the flow in the trans-sonic region. He has shown that the equation which describes the problem has two complex roots in addition to the root discussed by Nelson and Hearn, and that these roots correspond to strongly amplified sound waves. Martens suggests that the amplified sound waves which come from his solutions may be the source of macroturbulent velocities of the order of the local velocity of sound in the layers of the atmosphere near the sonic point. We have noted above the evidence in the spectra of τ Sco, 10 Lac, and ζ Pup for large macroturbulent velocities; Ebbets (1979) has noted evidence for the presence of macroturbulence of the order of 30 km s^{-1} in the atmospheres of OB supergiants. Martens has used simple models for the atmospheres of ζ Pup and ϵ Ori, B0 Ia, to estimate how much energy might be deposited in the trans-sonic region by amplified sound waves, and he has indicated that enough energy may be deposited to produce the hèated regions which are inferred to be present in the mantles of ζ Pup and ϵ Ori.

B. Stability in the Supersonic Region of the Wind

MacGregor et al. (1979), Carlberg (1980), and Kahn (1981) have studied the stability of radial outflow in the supersonic part of a wind which is driven by a force which is proportional to the velocity gradient. Such a force has been deduced by Castor (1974b) to result from the radiation absorbed in lines (see the section on the *Mechanical Force Exerted by Radiation*). The above authors show that laminar radial flow is unstable to perturbations in the temperature, the density, and the speed of outflow. MacGregor et al. indicate that the dissipation of the amplified perturbations (sound waves) may be an important source for heating the flow, particularly in the supersonic region. Carlberg shows that the radiation-driven sound-wave and gradient instabilities which may occur have properties such that the wind will tend to break up into clouds or slabs having typical sizes from 10^9 to 10^{11} cm. Such blobs may account in a qualitative manner for the discrete components which are seen in the winds from O stars (Lamers, 1981). Typically, the radius of an O star is 15 to 20 $R_\odot$ (3×10^{11} to 4×10^{11} cm). Consequently, the blobs have sizes which are small with respect to the disk of the star. It is probable that two or three blobs may be seen projected against the disk of the star at any one instant. The presence of such blobs or clumps of material would alter the dynamics of a wind and the interpretation of the profiles of lines formed in a wind from the picture which has been developed using the geometric description of spherically symmetric laminar outflow. A detailed theory which accounts for the discrete components has been developed by Underhill and Fahey (1984). They postulate the release of parcels of gas from a rotating star which has bipolar magnetic regions on its surface.

C. Generation of X Rays

Cassinelli et al. (1978) and Cassinelli and Olson (1979) have argued that a source of X rays concentrated in a slab at the base of the wind from a luminous O or B star may provide, by Auger ionization, the high ions which are seen by means of absorption lines formed in the winds from luminous O and B stars. However, these authors do not indicate the manner in which the postulated X rays are to be generated. Presumably, the X rays are to be the result of thermal radiation from parcels of gas having temperatures near 10^7 K; no heating mechanism is specified in the quoted papers. Recently, Cassinelli et al. (1981) have noted that a slab of very hot material at the base of the wind cannot be the source of the X rays which are detected from the luminous O and B stars. This conclusion is reached because the rates of mass loss deduced for the luminous O and B stars from observed and inferred radio fluxes imply a column density of material in front of the star which will attenuate X rays formed at the base of the wind by more than what appears to be the case for the observed X rays. However, Waldron (1984) has demonstrated that X-ray spectra like what are observed may be predicted if one models the wind in detail.

Cassinelli et al. (1981) have concluded that if one wishes to account for the observed X rays from luminous O and B stars, one is restricted to two types of model: (1) a model in which the observed X rays are formed in hot material distributed throughout the wind, and (2) a model in which some deep-lying very hot volumes of gas are not overlaid by the relatively cool gas of a wind which is in radiative equilibrium at a temperature close to the effective temperature of the star. The latter configuration may result from the modulation of surface magnetic fields by conditions in the stellar interior and from the stressing of surface magnetic fields by the rotation of the star (see Vaiana et al., 1981).

Pallavicini et al. (1981) have shown that, in the case of the luminous O and early B stars, the observed X-ray luminosity correlates well with the bolometric luminosity of the star. This finding is readily understood in terms of the picture developed earlier in this chapter and in Chapters 4 and 8 of Underhill and Doazan (1982). This picture postulates that conditions in the mantles of luminous O and B stars may be the result of the presence of small magnetic fields which have erupted from the photospheres of the luminous early-type stars.

Lucy and White (1980) have proposed that instabilities in the winds from hot stars, such as those discussed in the previous section, create blobs of material and that these blobs develop hot bow shocks as they are driven by radiation absorbed in the resonance lines of high ions through the relatively cool part of the wind. This cool part of the wind is flowing outward at lesser speed than that at which the blobs flow. The very hot parcels of gas in the bow shocks then radiate X rays. Acceleration of the fast outward motion of any particular blob will cease after a short time as the blob is shadowed by other blobs closer to the star and the blob suffers a diminished radiative force. The blob will cool by emitting radiation, and eventually it will merge with the general wind. Because many blobs may be expected to be present around a hot star at any one moment, the flux of X rays observed from the hot star is expected to remain approximately constant.

The theory of Lucy and White is not entirely satisfactory because it predicts greater attenuation of the X-ray flux at K-shell edges of C and O than the observations imply. Lucy (1982) has modified the theory to include a mechanism for dissipating the radiatively driven shock fronts more rapidly than was inferred by Lucy and White to occur. The revised theory predicts that X-ray emission continues far out into the terminal flow. Consequently, the escaping flux of X rays suffers little self-absorption with the result that the predicted spectrum shows less prominent absorption edges than those found by Lucy and White. The shape of the predicted X-ray spectrum is now more like what is observed, but Waldron (1984) has noted other problems with the theory.

It is clear that plasma instabilities of mechanical and, possibly, of magnetohydrodynamic origin are important sources of energy for heating the winds from O and Wolf-Rayet stars. They are also important sources for creating an inhomogeneous geometric structure. Models of mantles which are more sophisticated and more realistic than those reviewed in this chapter will have to be studied before we can say that we understand how the mantles of hot stars are created and how the mantles obtain their properties of high electron temperature in some regions, outflow, and inhomogeneity.

8

THE PHYSICS OF THE MANTLES OF HOT STARS

I. THE SITUATION EXISTING AT THE END OF 1983

In the preceding chapter, the results published chiefly before the end of 1983 relevant to spectrum formation in the atmospheres of O and Wolf-Rayet stars have been reviewed. We have seen in the first third of Chapter 7 that spectroscopic features which originate in the photospheres of O and Wolf-Rayet stars can be understood in terms of traditional theory. However, we have seen also that conspicuous parts of the spectrum such as a continuum of infrared and radio-wavelength excess energy, the profiles of leading members of sèries of lines from abundant atoms and ions, strong emission lines, especially those from ions requiring much energy for their formation, and X rays cannot be explained by traditional theory. The profiles of resonance lines in the ultraviolet spectra of O and Wolf-Rayet stars have revealed outflow at speeds exceeding the velocity of escape from the photosphere of the star. In the case of some stars, moderately strong, shortward displaced, sharp absorption components of the resonance lines are present. It has been argued that these lines imply the action of local magnetic fields (Underhill and Fahey, 1984; Mullan, 1984). We suggested in Chapter 6 that the part of the atmosphere in which those parts of the spectrum which cannot be explained by traditional theory are formed should be called the mantle of the star.

In the last two-thirds of Chapter 7, we have reviewed the published theories dealing with the nontraditional parts of the atmospheres of O and Wolf-Rayet stars. In the section *The Line Spectrum from Moving Three-Dimensional Atmospheres*, we showed that laminar spherically symmetrical outflow from a hot star gives an acceptable explanation for some, but not all, of the details of the absorption and emission lines formed in the mantles of O and Wolf-Rayet stars. However, in the section *Atmospheric Stability: X Rays*, we argued that laminar flow explains neither the presence of discrete, rather sharp, displaced components of the ultraviolet resonance lines of O stars, nor the changes of these components in times of the order of hours to months and their longevity in the case of supergiants, nor the facts that ions requiring much energy for their formation, such as O^{+5}, are seen in the atmospheres of relatively cool O stars, and that X rays are emitted by O and Wolf-Rayet stars in small amounts.

We have noted the possible importance of magnetic fields and the dissipation of magnetohydrodynamic (MHD) waves for accounting for what is seen (see the section *The Physics of Winds*), and we have concluded that the fraction of the stellar surface covered by magnetic flux tubes appears to be a significant parameter for determining the state of the mantle, thus of determining the major properties of the spectrum which is seen. We have noted also the substantial "turbulence" at supersonic velocities which appears to be present in the mantles of O and Wolf-Rayet stars, and we have seen in the section *Atmospheric Stability: X Rays* that plasma instabilities may cause part of what we see.

In the section *Models for Radiatively Driven Winds Powered by Absorption in Lines*, it was noted that the winds of the early-type stars may be like the winds of late-type stars in that their origin appears to be a result of the action of magnetohydrodynamic forces. The winds of early-type stars differ from the winds of late-type stars in that, in the winds of hot stars, the particles may be accelerated to high velocities by means of the force of radiation in addition to any MHD forces which may be active. In the case of late-type stars, acceleration by the force of radiation coming from the photosphere is negligible unless dust is present. It was concluded that stars of the same spectral type may be expected to have similar winds only if they have similar distributions of surface magnetic field.

II. MODELING PROCEDURES FOR MANTLES

We have seen that all of the features which give the spectra of O, Of, and Wolf-Rayet stars their particular character are formed in the mantle of the star. Consequently, to understand what O and Wolf-Rayet spectra mean, we must develop models for the mantles of O and Wolf-Rayet stars. First, we shall look at the problem in broad outline, and then we shall review what is indicated by some attempts to make generalized models for mantles. As a very minimum achievement, we hope to determine from generalized models what factors are important for creating certain sets of observable conditions and which are of little importance. What we deduce for the Population I O and Wolf-Rayet stars may also be relevant for understanding the spectra of hot subluminous stars. We conclude Chapter 8 by reviewing how the presence of magnetic fields may help us to understand the meaning of the spectra of O and Wolf-Rayet stars.

A. The Problem in Broad Outline

All modeling is begun by considering the equations describing the conservation of mass, momentum, and energy and applying constraints on what fluxes of mass, momentum, and energy may exist at each point in the model. These constraints are applied in order to reduce the general conservation equations to forms which describe the particular case under study. In the work to be discussed below, little account is taken of possible changes with time of the parameters which describe the state of the system, either on short time scales of a few seconds in length or on long time scales of a few years. A steady state is assumed to exist. The equation of state is that for a perfect gas having an assigned composition.

In the case of all massive stars, a flux of radiation, generated by nuclear reactions in the core of the star is postulated to flow through the body of the star and the stellar atmosphere. In the present discussion, we shall not consider pulsations which may be generated in the envelope of the star. Rather, we shall adopt the point of view that the pressure structure of the model star is such that each layer of the envelope and photosphere is in hydrostatic equilibrium under the inward-directed force of gravity due to the mass of the star and the outward-directed force of radiation. We postulate that no flows occur, except perhaps the cyclic motions of convection in some deep zones of the model.

Parker (1981), working from fundamental principles, has estimated the consequences for heating the outer solar atmosphere of imposing a spherically symmetric flow at the photospheric level, whose mass flux equals that observed in the solar wind by satellites at 1 a.u. He concludes that the flow will not achieve sufficient velocities where the solar chromosphere and inner corona are located to produce heating in the amount required to produce observed chromospheric and coronal temperatures. On this basis, he questions the efficiency of a photospherically imposed velocity field to produce, ultimately, the solar wind, if this wind is due primarily to the expansion of a heated corona by the thermal pressure gradient, as originally proposed by himself. Parker's comments were made for solar-like, late-type main sequence stars. However, we might also consider them in the context of winds from hot stars, because they focus on the action of energy and forces in the mantle usually ignored when the winds of hot stars are modeled, as for instance by Castor et al. (1975), Cassinelli and Olson (1979), Cassinelli et al. (1981), and Waldron (1984).

Traditionally, models of massive stars are constructed with the constraint that, outside of the core, the total amount of energy in the radiation field is conserved (i.e., that radiative equilibrium occurs) and that no additional source of energy is significant anywhere in the model. The constraint that hydrostatic equilibrium should occur throughout the model is applied also. Consequently, it is postulated that there is no systematic transfer of mass and momentum from one layer to another anywhere in the model. These constraints serve to define acceptable models for the interiors and photospheres of massive stars.

It is recognized that, at some depths in the envelope of a massive star, convection may be important for transporting energy; also, convection is believed to occur in the core of a massive star. Nowhere in the photosphere and interior of the star is conduction an important mechanism by which to transport energy. It may, however, be significant when one is modeling the mantle of a star. Also when modeling the mantle of a star, one must consider the transport of energy by the flow of particles and as a result of the changing internal energy of the transported material.

Because most of the radiative energy observed to be emerging from a star is generated by nuclear reactions in the center of the star, one has to take account of the flow of radiation throughout a model star. At present, however, there is no known physical mechanism indicating that the observed outflow of mass from stars of any spectral type is due to a flow originating deep in the star. Models based on the postulate that no mass and momentum flows occur below the photosphere of the star are adequate to explain many of the observed properties of almost all stars.

It is not necessarily true that the material seen to be flowing from the surface of a star is part of a recognizable flow originating deep in the star. The release of material into a wind may be a surface phenomenon. The star, of course, will readjust its pressure structure to take account of the loss of mass, but this action produces negligible hydrodynamic flow because the fraction of the star's mass which is lost in this way is very small during an interval of time typical for movement at the speed of any conceivable flow.

The significance of Parker's (1981) study is his claim that coronal heating, which can produce some of the observed solar wind, cannot result from the normal hydrodynamic expansion of the small photospheric velocities required to balance solar wind mass losses. It is known, on the other hand, that much cyclic flow arranged in cells of different sizes (granules and supergranules) occurs in the deep layers of the solar photosphere. Parker argues that any outward-directed flow that might be generated from these cyclic motions is too small to affect the solar wind in the manner advocated by Thomas (see Chapter 7, *Characteristic Features of the Stellar Wind Problem*).

It is true that modern observations have shown that a small amount of mass flows from most, if not all, massive stars at velocities exceeding the velocity of escape from the star. Any successful model of the mantle of a massive star should take this observed fact into account. Part of the problem of making acceptable models for mantles concerns determining where the observed outward mass and momentum flows are generated, how the material is accelerated outward, and how much material is in the flow.

It is only in the mantle of the Sun (i.e., in the chromosphere and corona) that part of the gas is *observed* to acquire systematic outflow velocities. Disturbances called *coronal transients* occur from time to time, and prominences sometimes erupt. A low-density wind of nonuniform structure, observed by means of spacecraft and from a study of the behavior of comet tails, is known to be flowing in the region around the Sun. There is observational evidence suggesting that material is released from gravitational and magnetic control near the surface of the Sun at particular spots and is propelled outward, sometimes in the form of impulsive events and sometimes as a general rather uniform wind.

On the foregoing basis, we believe that it is appropriate to consider the possibility that the observed outflow is generated at the interface between the photosphere and the mantle, in the way that many solar physicists currently believe occurs for the Sun.

The detailed theories which deal with radiatively driven winds all begin with the ad hoc assumption that a gradient of flow velocity already exists in the mantle of the star. Any acceptable theory of the winds from hot stars must include a way of starting the wind and of setting up velocity gradients in the mantle.

No matter how small the mass flux may be at any point in a star, that mass flux must be generated by the action of some net unbalanced force. No one has identified the particular forces for generating a small flow deep in a star. On the other hand, there is much evidence from the Sun that MHD forces act in the mantle of the Sun to initiate outflow either in the form of large parcels of material released from particular spots above the surface of the Sun (coronal transients) or as a low-density, rather uniform wind emanating chiefly from coronal holes. Parker (1981) argues that the pressure gradient resulting from the high temperature of the solar corona is enough to initiate outflow of the mean solar wind. The high temperature is believed by some solar physicists to result from dissipation of mechanical energy in the presence of magnetic fields (Wentzel, 1981).

We shall explore the applicability to hot stars of components of the picture used to account for the physical state of the corona and wind of the Sun. In his review of solar conditions, Wentzel (1981) has emphasized that, although the active corona of the Sun is the site of primarily closed magnetic fields, the high-speed solar wind emerges from regions of the corona which consist of primarily open magnetic fields (i.e., of field lines which extend far into interplanetary space). The presence of the solar wind and of the magnetic fields which it transports can be inferred at the distance of Saturn from studies of the magnetosphere of Saturn (McDonald et al., 1980; Vogt et al., 1981), while observations made from Pioneer 10 and 11 (McDonald et al., 1981; Collard et al., 1982) have revealed the character of the solar wind and its associated magnetic fields at distances from the Sun greater than 4500 $R_\odot$. It should be noted that the electron temperature of the solar wind is observed to be high at these distances and the material has a high outflow velocity, a considerable amount of acceleration having occurred from the low flow velocities inferred by Rottman et al. (1982) and Orrall et al. (1983) to be present near the base of the corona.

Conditions observed by means of spacecraft to exist in interplanetary space present examples of the type of situation which we may expect to arise in the vicinity of O and Wolf-Rayet stars. The most significant differences between the physical situation near hot stars and that which exists near cool stars appear to be the differences in the density and energy distribution of the radiation field from the star and in the density of the wind.

In the case of O and Wolf-Rayet stars, we have to infer the conditions in the wind from modifications impressed on the stellar spectrum as the light of the star traverses the stellar wind. In the case of the Sun, it is possible to sample conditions in the wind by making in situ measurements. Since the *physics* of a mixture of gas, radiation, and magnetic and gravity fields must remain the same for stars as for the Sun, it is plausible to expect that observations of what happens near the Sun may give insight into what phenomena may be occurring near stars. The particle and radiation densities in the winds from hot stars are believed to be greater than those in the solar wind, to be sure, but we believe this is unlikely to change the *physics* of the problem in a fundamental way.

One reason why it is desirable to consider the action of magnetic fields in the mantles of O and Wolf-Rayet stars is that the spectra of O and Wolf-Rayet stars give evidence for the presence of plasma heated to electron temperatures exceeding the effective temperatures of O and Wolf-Rayet stars by a large amount. We recall that the effective temperatures of Wolf-Rayet stars are of the order of 25000 to 30000 K, while those of O stars lie in the range from 30000 to 45000 K in most cases (Underhill et al., 1979; Underhill, 1980a, 1981, 1982, 1983b; and Part I of this book). A second reason is the common presence of discrete absorption components superposed on the "wind" profiles of hot stars (see the discussion by Underhill and Fahey, 1984). A third reason is the intrinsic spread in the terminal velocities seen for stars which have comparable effective temperatures and luminosities, and thus, presumably comparable masses and radii. Underhill (1983c) has suggested that this spread indicates that one of the accelerating forces acting in the wind of a hot star is that due to the presence of Alfvén waves. A fourth reason is the surprisingly large radio fluxes received from Wolf-Rayet stars. Underhill (1983a, 1984a) has argued that most of the radio flux comes from plasma suspended in magnetically supported loops.

Rosner et al. (1978a, 1978b) have argued that the solar corona is heated by the dissipation of mechanical energy in the presence of magnetic fields, and much observational evidence supports this point of view (see, for instance, Wentzel, 1981). The arguments used to justify the consideration of the effects of magnetic fields in the case of the Sun may be valid also for the mantles of O and Wolf-Rayet stars.

The temperatures believed to occur in the mantles of O and Wolf-Rayet stars are sufficiently high that heating by the dissipation of rotational energy as a result of interactions with a magnetic field is an attractive possibility to consider. Heating resulting from the dissipation of compressional waves generated in the convective zones of O and Wolf-Rayet stars is a less attractive possibility because the subphotospheric convection zones of hot stars are not expected to be as large relative to the size of the star as is the case for the Sun. The dissipation of compressional waves is not adequate for heating the solar corona (see Rosner et al., 1978a, 1978b; Wentzel, 1981), although it may be significant for heating the solar chromosphere. That the classical assumptions (Parker, 1958) do not give a satisfactory interpretation of the solar wind has been noted by Hollweg (1981) and Kopp (1981). They review what nonthermal processes may be significant.

In what follows, we shall assume that some of the magnetodynamic phenomena known to occur in the Sun are occurring in the atmospheres of O and Wolf-Rayet stars, and we shall try to assess the importance of such phenomena for accounting for features in the spectra of O and Wolf-Rayet stars which cannot be explained by traditional models of atmospheres and the process of forming a spectrum.

It is generally conceded that O stars rotate rather rapidly (see, for instance, the observations reported by Wolff et al., 1982). In addition, spectroscopic analysis has indicated that a significant amount of turbulence is present in the photospheres of supergiant O stars (Ebbets, 1979). This means that there is differential motion in the photosphere, perhaps overturning motions connected with the formation of vortices. If such motions occurred in the presence of weak primordial magnetic fields which had been occluded as the star formed, local dynamos would be created. One may expect that many bipolar magnetic regions will be formed in and above the photospheres of O stars. The probability of this being so is sufficiently great as to make the presence of weak local magnetic fields in the atmospheres of all O stars very likely. Suggestions have been made that ζ Pup, O4ef, has a small magnetic field (Mihalas and Conti, 1980; Moffat and Michaud, 1981), but observational evidence for a magnetic field in ζ Pup is inconclusive (Barker et al., 1981).

Because the emission lines formed in the mantles of Wolf-Rayet stars mask almost all of the spectrum from the photosphere of a Wolf-Rayet star, it has been impossible to demonstrate by direct measurement that magnetic fields are present in the photospheres of Wolf-Rayet stars. However, the evident conspicuous superheating of the mantles of Wolf-Rayet stars (see Chapter 7, *Theories for Interpreting Wolf-Rayet Spectra*) and the need to account for the observed radio fluxes from Wolf-Rayet stars without postulating an unphysical situation (Underhill 1983a, 1984a) provide strong circumstantial evidence that magnetic fields are present in the photospheres of Wolf-Rayet stars. Nussbaumer (1982) has commented that magnetic flux may be an important agent for producing winds from late-type stars and from the central stars of planetary nebulae. Most central stars of planetary nebulae have O or Wolf-Rayet spectral types.

B. General Methods for Modeling Mantles

The distinctive features of the coronas of cool dwarf main-sequence stars are: (1) strong gravitational binding, the gravitational binding energy being five or more times the thermal energy, and (2) direct deposition of energy to produce a high temperature. Parker (1981) has concluded that the mean motion of the photospheric gas may supply mass to the stellar wind, but that it will not supply a significant amount of momentum to the wind.

In the case of a large cool star with low surface gravity, thus a weakly bound corona, or in the case of a star with a dense radiation field which reduces the effective value of gravity in the corona to small or negative values by producing a significant value for g_{rad} in the corona, the situation is more complex than it is for cool dwarf stars. Parker (1981) states that this complex type of problem cannot be handled by a simple analytical treatment.

The response of the outer atmosphere of a cool low-gravity star to the passage of mechanical energy fluxes in the form of acoustic waves and of Alfvén waves has been studied by Hartmann and MacGregor (1980). These authors find that acoustic waves do not seem to be important for driving mass loss, but that Alfvén waves may be. The Alfvén waves must be dissipated by an appropriate amount low in the corona of a giant cool star if winds having low terminal velocities, as observed, are to be generated.

Simplified models of regions in the hot coronae of cool stars have been constructed by Hammer (1982a, 1982b, 1984). Hammer has developed these models for the purpose of understanding systematic trends among the properties of the mantles of F-, G-, and K-type giants and of cool main-sequence stars. The consequences of his theory are of interest for understanding the mantles of early-type stars as well. The chief difference between the mantles of the stars studied by Hammer and those of interest in this book is that, in the case of the late-type stars, the density of the radiation field from the star at wavelengths shorter than about 2000 Å is negligible, while in the case of the early-type stars, there is much radiation at wavelengths shortward of 2000 Å passing through the mantle. The radiation which is present in the case of hot stars may interact with material in the mantle and affect the physical state of the mantle. The problem of generating model mantles for the early-type stars might be approached by starting with the work of Hammer, modifying, as appropriate, the treatment of mechanical heating, and adding the effects of radiation on the wind. This will be a formidable numerical problem but one which must eventually be tackled as our understanding of the underlying physical mechanisms improves.

The model mantles developed by Hammer have spherical symmetry, and they are composed of fully ionized hydrogen. Hammer notes that, at the time he wrote his papers (1981–1982), the mechanism which causes coronal heating was unknown, although some plausible suggestions had been made. The physical state of the model corona depends sensitively on the character of the heating mechanism and on where in the corona the heat is deposited. In the section *The Problem in Broad Outline*, we have reported the arguments which have been made linking coronal heating to the presence of magnetic fields in the corona.

Hammer develops his models in terms of a parameter, ϕ_M, which tells how much "mechanical" (i.e., nonradiative) energy is deposited in the corona, and the damping length, L, which tells where in the corona the nonradiative energy is dissipated. He assumes an exponential law to represent the dissipation as a function of height in the corona (mantle), but also explores the effect of a more complicated expression which may describe heating resulting from the dissipation of weak shock waves. The results are not very sensitive to the form adopted for the heating law. Two other parameters are important for defining the model mantles. These are the stellar radius and mass. These quantities define the size of the base of the mantle and the potential field, due to gravity, in which the material of the model corona moves.

In the case of a hot star, one can picture that radiation pressure reduces the value of gravity to a small positive value or even to a negative value. The passage of a field of radiation through a low-density hot gas does not heat the gas to temperatures exceeding T_{eff} in either LTE or non-LTE. Thus, the temperature in a stationary hot spherical atmosphere which is in radiative equilibrium decreases outward steadily, but not monotonically. In my opinion, there is no reason to believe that flow of the gas in the spherical atmosphere will change this result beyond causing localized heating in a few places as a result of the dissipation of energy in instabilities arising in a flowing gas (see Chapter 7, *Atmospheric Instability: X Rays*).

Hammer derives forms of the conservation equations appropriate for the case that mass flows outward with a velocity which is dependent only on radius. Mass is conserved in concentric spherical shells as it flows. Hammer considers four ways in which energy may be transported; namely: (1) by a wind energy flux which is given as the sum of the kinetic energy of the mass flowing in the wind, the enthalpy, and the potential energy of the material, (2) by a conductive energy flux which depends on the composition of the material and the temperature gradient existing at each point in the model mantle, (3) by the loss of radiative energy owing to the escape of photons from the mantle, and (4) by the flux of "mechanical" energy which is postulated to be deposited in the mantle. The "mechanical-energy" flux postulated by Hammer includes any net deposit of energy which may occur owing to MHD interactions. It is described by an exponential function which depends on two assigned parameters—the mechanical-energy flux entering the base of the model mantle, ϕ_{M_0}, and the damping length, L.

The three ordinary differential equations of first order for the three variables (flow velocity, isothermal sound speed or temperature, and net flux of energy lost from the star) are solved subject to assigned conditions at the base of the corona, at the critical point where the flow speed equals the isothermal speed of sound, and at infinity. It is assumed that each satisfactory solution must reach infinity with a finite flow velocity and with $T \to 0$ as $r \to \infty$. Hammer discusses relations, valid for the solar atmosphere, which allow him to specify conditions at the base of the corona and at the critical point, and then to find appropriate solutions as r increases without limit. He varies the parameters of his models (ϕ_{M_0}, L, M_*, R_*) and obtains the variation with height above the base of the corona of the energy flux in the wind, that due to conduction, and that due to radiation, as well as the variation of temperature.

Hammer finds that three types of model mantle may exist for the types of star which he studies:

1. When the energy-dissipation scale lengths are small compared to the stellar radius, radiative losses dominate the energy losses from the mantle, compared with wind losses.

2. When the energy-dissipation scale lengths are large compared to the stellar radius and the heating fluxes are small, the outward conduction of heat contributes significantly to the energy losses, and the resulting coronal temperature is high; also, the base pressure is small, and the mass-loss rate is relatively small.

3. When the energy-dissipation scale lengths are large with respect to the stellar radius and large heating fluxes are present, the corona is wind-dominated, which means, here, that the radiative and conductive energy losses, as well as the base pressure, are small, while the coronal temperature is high and the mass-loss rate is large.

All of these results are for model stars with M_* and R_* equal to solar values. The rates of mass loss vary between about $10^{-15}\ M_\odot\ \mathrm{yr}^{-1}$ and $10^{-12}\ M_\odot\ \mathrm{yr}^{-1}$ as the amount of "mechanical" energy deposited is varied from about 5×10^5 to 2×10^6 erg cm^{-2} s^{-1}. The coronal temperatures are of the order of 10^6 K.

Making models is one thing; demonstrating which represents the mantle of a real star is another. Hearn (1975b) proposed that one may choose which is the correct model mantle for a real star by finding the minimum flux corona. However, detailed modeling such as that done by Hammer (1982a, 1982b), by Hammer et al. (1983), by Holzer and Leer (1980) and Leer and Holzer (1980), and by Holzer, Flå, and Leer (1983) has shown that Hearn's concept yields unsatisfactory results. It has been demonstrated by means of Hammer's models that the structure of the mantle and the size of the mass flux from the model, the maximum temperature that occurs and where it occurs, as well as the pressure at the base of the model mantle, depend very much on where the heat and momentum are introduced and on the path length over which the energy and momentum sources dissipate. The results also depend on how much energy and momentum are introduced at the base of the mantle.

Energy addition in the region of subsonic flow increases the mass flux, but it has little influence on the terminal speed reached by the wind. Energy addition in the region of supersonic flow increases the terminal velocity but has little effect on the mass flux (Kopp, 1981). Momentum transfer to the wind in the subsonic-flow region can increase the terminal velocity of the wind. Some momentum addition is needed to push the wind material through the critical point. However, the temperature gradient which exists in the neighborhood of the critical point is also important (see Equation (7-17) and also Equation (7-34), which was derived for the case of only radiative acceleration of the wind).

Hearn (1983) has used the concept of long-period mechanical waves in the mantle to simulate what happens when the deposition of the mechanical energy necessary to heat a mantle to coronal temperatures occurs over a long path length in a model mantle. Hearn's model mantle may be representative for an OB supergiant. He finds that a heating mechanism with a long dissipation length (cases 2 or 3 above) increases the fraction of the available mechanical energy which is taken up by mass loss at the expense of the fraction of mechanical energy which supports losses by radiation. Hearn's results indicate that energy loss through a wind will dominate in coronal-hole-type regions (in which magnetic lines of force extend long distances into the surrounding medium), whereas losses of energy by radiation will dominate when conditions are such as to confine the heating to relatively small regions of the mantle (case 1 above), as in a region dominated by the presence of magnetic loops. It remains to demonstrate what the actual mechanisms are for the deposition of energy which has originated as mechanical energy, but qualitatively, it seems possible that the different rates of dissipation of MHD waves in regions of open and closed lines of force may be a key factor in producing what is observed.

In Chapter 7, *Properties of Spherical Flow*, we have seen that radiation pressure is not an effective agent for initiating the mass flow in the wind from a hot star, although radiation pressure may accelerate flow once an adequate velocity gradient has been set up. Photoionization from ground levels is not an important process in the winds of normal hot stars because there is little radiation at the wavelengths (mostly shorter than 200 Å) of the chief ionization continua of the abundant ions. Because the material of the mantles of hot stars has a rather large opacity in resonance lines, and sometimes in a few lines which arise from low-lying metastable levels, we are able to see the material of the wind by means of absorption lines formed in the wind. This is not possible in the case of cool stars which do not have a hot companion.

It has been noted in Chapter 7, *Analysis of the Absorption Lines of O Stars: Abundances*, that many of the strong lines in the spectra of O stars are broadened, as though by macroturbulence. (The reader is cautioned that this phenomenon is called "microturbulence" in some of the relevant papers.) The typical velocity of the broadening motion appears to lie in the range from 100 to 150 km s^{-1}. It is possible that by means of the strong lines we are seeing parts of the mantle of the star which are essentially stationary so far as outflow is concerned, but in which the atoms and ions are taking part in gyro motion around magnetic lines of force in the mantle of the star. Such motion would simulate what is called "turbulence" when speaking of line-broadening factors. The moving electrons would radiate at radio frequencies. How significant gyro motion may be for generating the observed radio fluxes of hot stars has been discussed by Underhill (1984a). It appears that the rates of mass loss determined from observed radio fluxes and listed in Tables 7-9, 7-10, and 7-11 may be upper limits to what truly occurs. Reduction by a factor of the order of 10 is plausible.

The method of Hammer (1982a, 1982b) may be used to make numerical models of the mantles of O and Wolf-Rayet stars when this method is modified to include an equation for the flow velocity as a function of radius which is appropriate for a wind driven by radiation and by MHD waves. One must also introduce appropriate expressions telling where the heating occurs and from what places mass is released from gravitational control by the star. Is the wind released in a uniform spherical flow, or is it released preferentially from regions corresponding to the coronal holes on the Sun? In order to specify the needed expressions, one must make a model of the physical processes which occur in the mantle and determine the magnitude of the heating and outflow generated in each element of volume in the mantle. Thought must be addressed to the problem of describing how heating and the release of gas from gravitational control occurs in a plasma containing radiation and magnetic fields, as well as charged particles, and situated in a gravitational field.

In the remainder of Chapter 8, we explore the consequences of postulating that small magnetic fields are present in the mantles of O and Wolf-Rayet stars. The goal is to see if some of the problems with the existing theories of the mantles of O and Wolf-Rayet stars can be solved.

III. MAGNETIC FIELDS IN THE MANTLES OF HOT STARS

It has been pointed out by Underhill (1983a) and by Underhill and Fahey (1984) that, if one postulates that bipolar magnetic regions are present in the photospheres and mantles of O and B supergiants and of Wolf-Rayet stars, a natural explanation follows for many observed spectroscopic phenomena. As in the Sun, the average magnetic field over an element of surface having an area of the order of $\frac{1}{2} R_*^2$ may be small. In addition, the magnetic lines of force threading the photosphere may be concentrated into small areas of strong field—of the order of 1500 to 2000 gauss. This may be a result of the buffeting of the foot points of the magnetic lines of force by moving elements of plasma in the photosphere. The buffeting generates magnetohydrodynamic (MHD) waves in the volumes of plasma containing magnetic lines of force. The waves act on the plasma by heating it and by exerting a stress on the atmosphere. This stress may lead to an outward acceleration of any wind which may be flowing.

Because the theory of model atmospheres and predicted spectra (see Chapter 7) enables us to model the photospheric spectra of O stars rather well, we will propose here that most of the photosphere of a hot star is relatively free from magnetic lines of force, as in the Sun, and that the magnetic field lines which traverse the photosphere are concentrated into small volumes at the intersections between the moving elements of plasma which we will assume are generated by ambient turbulence. In the low-density plasma of the mantle, however, the structures containing magnetic lines of force will spread and grow, as occurs in the solar mantle. It is the interactions between radiation and plasma which occur in these large structures which give the spectra of O, Of, and Wolf-Rayet stars many of their dominant characteristic features.

A. General Remarks

First we shall review evidence suggesting that conditions favorable to the creation of magnetic flux exist in the photospheres of OB supergiants and Wolf-Rayet stars, and after that we shall note how the presence of magnetic lines of force in a mantle provides a basis for understanding the spectra of O and Wolf-Rayet stars. Magnetic lines of force act as conduits for the mechanical energy carried by the MHD waves generated by turbulent motions in the photosphere, closed loops acting like resonant cavities, open lines of force acting like wave guides (Davila, 1985). Magnetic lines of force can also store energy and momentum, and this energy and momentum can be released when magnetic flux is annihilated (Kuperus et al., 1981).

We postulate that magnetic flux is created in the deep parts of the stellar photosphere by some kind of suitable stellar dynamo. Dynamo action can result from overturning motion generated by the differential rotation of the star in the presence of primordial magnetic fields which were occluded as the star formed, and it is likely that the continuing turbulence in the photosphere will maintain this action. (See Cowling, 1981, for a review of dynamo theory.) It should be recalled in this context that the microturbulent velocity in the atmosphere of a main-sequence O star is typically about 5 km s^{-1}, while that in the atmosphere of an OB supergiant may be 15 km s^{-1}.

It may readily be demonstrated from model atmospheres which represent well the photospheres of stars having effective temperatures in the range from 25000 to 35000 K that the temperature gradient resulting from the assumption of radiative equilibrium exceeds the adiabatic temperature gradient in the layers in which the continuous spectrum in the visible range is formed. Consequently, conditions in these layers are appropriate for convection to become established (Underhill and Fahey, 1984).

When making model atmospheres for hot stars, it is not necessary to consider convection as a way of transporting energy through the upper layers of the envelope and through the photosphere because the transport of energy by radiation is very efficient. However, conditions making it possible for convection to take place may occur. This was discovered by Underhill (1950) using a simple model atmosphere. Little use has been made of this information because the model atmospheres found by assuming radiative equilibrium throughout the model atmosphere and no transport of energy by convection differ very little from what is found when an adiabatic temperature gradient is adopted in the layers in which convection may occur.

The temperature gradient, d(log T)/d(log P_g), in part of an NLTE model atmosphere by Mihalas (1972b) is shown in Figure 8-1, together with two estimates for the adiabatic temperature gradient in these layers. The adiabatic temperature gradient has been estimated using the expressions given by Underhill (1949). The results for the case which includes the effect of radiation (short dashes) probably are closer to what will exist in nature than are those for the case when the presence of radiation is ignored (long dashes). It is clear that at the depth where $m = 1$ g, the temperature gradient obtained by assuming radiative equilibrium throughout the model atmosphere begins to exceed the adiabatic temperature gradient. It remains larger than the adiabatic temperature as one goes deeper into the model. From Figure 8-1, we see that the layers in which the visible continuum

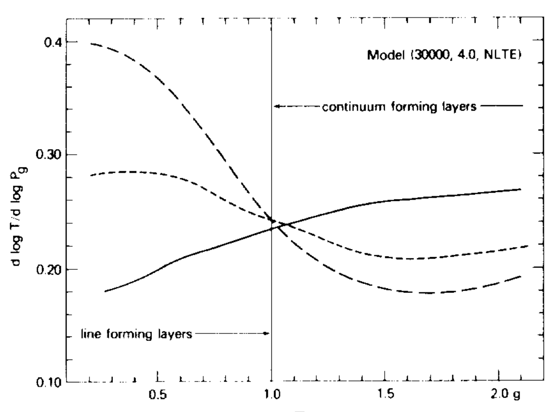


Figure 8-1. Temperature gradients, d(log T)/d(log P_g), as a function of depth in a model atmosphere. The gradient resulting from radiative equilibrium is shown by a solid line. The adiabatic gradient when the effects of radiation are ignored is shown by long dashes; that when the effects of radiation are taken into account by short dashes. In the continuum-forming parts of the model atmosphere, the adiabatic gradient is reduced from its value in the outer part of the model atmosphere as a result of the second ionization of helium. Depth in the model atmosphere is measured in terms of m, the mass above an area of 1 cm^2 in the model atmosphere.

is formed in the selected model atmosphere lie in the neighborhood of the depth point where $m = 1$ g. Consequently, we infer that convection may be established in the deeper layers of the photosphere. Differential rotation, added to this convection, could produce magnetic fields by a dynamo process similar to the one now believed to be operating for the Sun and some other late-type stars.

Because the layer in which He^+ is being ionized is not very thick in model atmospheres for hot stars, we do not expect the convective layer containing the postulated dynamos to be very thick. However, this point has not been fully explored. The linear scale of the convective layer in a star with small log g will be greater than that in a star with a relatively large value of log g.

Convective motion, differential rotation, and the presence of a weak primordial magnetic field are what is necessary for magnetic flux to be generated locally in the deep layers of the photospheres of stars which have effective temperatures of the order of 25000 to 35000 K. One expects, as in the Sun, that ropes of magnetic lines of force will be formed and that strands will erupt from the photosphere in places, generating bipolar magnetic regions (BMR's) on the surface of the star. We expect that the bipolar magnetic regions will be concentrated toward the equator of the rotating star, just as is true for the Sun. However, we are unable at present to confirm these expectations with a quantitative description of what may occur in a model hot star because too little is known at this time to enable one to apply dynamo theory to the hot stars.

The presence of magnetic lines of force in a model atmosphere requires us to revise the way in which the conservation equations (Equations (7-13), (7-14), and (7-15)) are constrained to represent reality. Not only does a magnetic field exert a stress perpendicular to the direction of the magnetic field, but the mobility of ions and electrons is different perpendicular to the field lines than it is in the direction of the field lines. In addition, the speed at which MHD waves propagate through the atmosphere depends on the magnitude of the magnetic field, a local quantity. The rate of dissipation of the energy in the MHD waves is field-dependent (see, for instance, Ionson, 1982, 1983; Rae and Roberts, 1982; Heyvaerts and Priest, 1983), and it depends on the local microscopic properties of the plasma. Annihilation of magnetic flux will occur locally as neutral lines are formed between areas of opposing magnetic polarity (see Parker, 1983a, 1983b for a summary of what may occur).

Interactions between ions, electrons, magnetic fields, and waves are observed in the Sun, but no fully accepted theoretical description of what occurs microscopically is presently at hand. Discussion is given below of some points which become important when one admits that magnetic lines of force are present locally in the mantles of O, Of, and Wolf-Rayet stars, and one postulates that the mantles of O, Of, and Wolf-Rayet stars consist of arcades of magnetically supported loops, as well as of a wind which contains lines of force open to the interstellar medium.

B. The Specific Problem of the Mantles of Hot Stars

We shall now review how a model based on what has been observed to be true for the Sun may serve to provide an understanding for many of the phenomena seen in the spectra of early-type stars from X rays to the radio range. First let us describe the model.

We shall think of the mantle of an early-type star as being composed of arcades of coronal loops extending on the average to a distance of the order of 20 R_* and of areas covered by structures equivalent to the coronal holes of the Sun. We expect the loop structures to occur predominantly in an equatorial band which extends, perhaps, to latitudes of the order of ± 30 degrees. The magnetic field lines emerging from the coronal-hole-type areas will emerge radially from the star and extend to great distances from the star. Here, in analogy with the Sun, we expect a high-speed wind to flow.

It is postulated that the magnetic lines of force supporting the arcades of loops are arranged with their foot points in BMR's. As in the case of the Sun, we expect the BMR's to be continually shifting position and changing in strength as more magnetic flux emerges from the photosphere and some magnetic flux is annihilated. Underhill and Fahey (1984) have summarized the reasons why such a picture is reasonable. The star rotates.

The magnetic field lines leaving the star from coronal-hole-type areas may have positive or negative polarity; they are traced by the particles flowing from the holes in the form of a wind. We expect that the magnetic field averaged over the full disk of the star will be small, as is the case for the Sun. The average field in one polar cap may have opposite polarity to that in the other polar cap.

We do not know how much, if any, of the surface of the star is free of magnetic lines of force. In the case of the Sun, fairly large areas of the photosphere are dappled with both positive and negative magnetic fields. Such areas, when observed from a great distance, can appear to be field-free. Magnetoplasmas originating above BMR's and forming part of the mantle may spread out over the apparently field-free regions of the photosphere.

In the case of the Sun, the magnetoplasma carrying the magnetic lines of force emerging from solar coronal holes appears to extend outward to about the heliopause, some 10^4 solar radii distant. Observations carried out from spacecraft (Burlaga et al., 1983; Burlaga, 1983) indicate that the spatial structure of the solar wind may become quite tangled before the wind merges with the interstellar medium. We expect that the coronal-hole-type regions on hot stars will behave like solar coronal holes and that the individual streams of the stellar wind emerging from coronal-hole-type regions will become tangled and merge with the interstellar medium under shock conditions far from the star. How uniform the stellar wind appears to a distant observer will depend on how many coronal-hole-type regions there are, on how they are distributed on the surface of the star, and on how unchanging the distribution of such regions is.

We expect the model which is described above to be valid for Wolf-Rayet stars, normal O and B stars, and Be, Oe, and Of stars. We exclude the Bp stars from our discussion. These stars have measurable longitudinal magnetic fields. Although it is conceivable that many of the ideas which we shall review here may be applicable to the Bp stars, we do not wish to discuss the Bp stars. The presence of sizeable magnetic fields traversing significant areas of the photosphere (spots) may well change the average conditions in the photosphere from what analysis of the spectra of normal nonmagnetic stars suggests is true. Also, sizeable magnetic fields will strongly constrain the motions possible in the atmosphere of the star.

When one is considering the events which may take place in the mantle relatively near to the surface of the star, say within about 20 R_*, it is essential, according to our model, to take account of the inhomogeneous structure of the mantle. When one is concerned only with the typical magnetic-field structure which may appear to a distant observer to be present, it may be advantageous to represent the average magnetic field by a potential function such as the field generated by a central dipole

I choose to explore the results which may be reached by considering the inhomogeneities of the mantle because it appears that the heating of the plasma in a stellar mantle and the release and propulsion of a wind may be events which take place in locations which are small in comparison to the full extent of the mantle and the size of the disk of the star. After an adequate description of the physics of what occurs at each place has been obtained, it may be possible to set up and solve the equations of radiative transfer and predict the spectrum which a distant observer will see.

The method which has been developed for calculating the spectrum from a nonmagnetic early-type star is to use the relevant physics of the problem under consideration to define the source function at each radial distance from the surface of the star, and then to solve the equations of radiative transfer appropriate to slab or spherical geometry and to predict what an observer will see. The results of such studies for nonmagnetic stars have been reviewed in Chapter 7. Symmetry with respect to the angular coordinates is assumed.

It may prove advantageous when dealing with that part of the mantle of an early-type star which is inhomogeneous (that part where arcades of magnetically supported loops occur and where coronal-hole-type structures exist) to define the source function in terms of a statistical description of what the physical situation may be at each point in the annulus where the loops predominate and in each polar cap. Once a description of the source function has been obtained as a function of the geometrical coordinates, it is, in principle, straightforward to set up the equations of radiative transfer and to estimate what sort of spectrum should be seen by a distant observer.

The first task is to see what can be said about the factors which determine the source function in various wavelengths as a function of position in the mantle. We shall need to know the electron temperature, T, the magnetic field, B, and the ion and electron densities, N_i and N_e, as functions of position in the mantle. We assume that a flux of radiation from the photosphere flows through the mantle and that additional photons are created or annihilated according to the physics of a moving ionized gas in the presence of radiation and a magnetic field. The conservation equations for energy, momentum, and mass provide a starting point. The full task is formidable. We shall begin by discussing the chief factors which should be considered. The equation of state is that of a perfect gas with solar composition.

C. Sources of Heat and Momentum in a Mantle Containing Magnetic Fields

When the plasma of a stellar mantle is laced with magnetic lines of force, the physics which must be considered in order to obtain a reasonably accurate representation of what occurs in the plasma may be selected by considering the value of the quantity β which measures the gas pressure due to electrons and heavy particles in terms of the magnetic pressure. The plasma β is defined by the relation,

$$\beta = 8\pi P_g/B^2, \tag{8-1}$$

where B is the magnetic field in gauss, and P_g is expressed in dyn cm^{-2}. In a plasma such that

$$\gamma(N+N_e)/2N \sim 1, \tag{8-2}$$

where $\gamma = C_p/C_v$ is the ratio of specific heats, and N is the number density of heavy particles in all stages of ionization, one can write

$$\beta \sim a^2/v_A{}^2. \tag{8-3}$$

Here a is the velocity of sound in the plasma, and v_A is the Alfvén velocity. Frequently, the β of stellar atmospheres is estimated by means of Equation (8-3) rather than by Equation (8-1). At most, β will be overestimated in a fully ionized cosmic plasma by a factor of the order of 2.5 when Equation (8-3) is used in place of Equation (8-1).

The significance of the plasma β is that when $\beta >> 1$, gas pressure dominates over magnetic pressure, and the effects of the magnetic pressure may be ignored when determining the hydrodynamic state of the plasma. When $\beta \sim 1$ or $\beta < 1$, the magnetic pressure is comparable to or greater than the gas pressure. Then the physics of magnetohydrodynamics must be employed.

The photospheres of stars may usually be treated as "high-beta" plasmas except for the volumes where concentrated magnetic lines of force pass through the photosphere. On the other hand, the mantles, like the small parts of the photosphere containing concentrated magnetic lines of force, require the use of MHD physics because in these places one has a "low-beta" plasma.

In the photosphere of an early-type star, the gas pressure may be typically about 700 dyn cm^{-2}. The plasma β will be 1, 2, 5, 10, or 20, respectively, as the magnetic field is 133, 94, 59, 42, or 30 gauss. We see that, if the average magnetic field in each element of volume is less than 100 gauss, we will have a "high-beta" plasma. We infer that the effects of magnetic fields will be negligible in all parts of the photospheres of early-type stars except those places where the lines of force are strongly concentrated.

In the mantle of an early-type star, however, the gas pressure may be of the order of 3 dyn cm^{-2} low in the mantle at $R = 1.25\ R_*$ and near 0.7 dyn cm^{-2} at about 5 R_*. In those parts of the mantle where $B \geqslant 10$ gauss, β will be less than 1. It is clear that MHD effects must be considered in all parts of the mantle in which magnetic fields greater than about 10 gauss are suspected to be present. In the parts of the stellar atmosphere in which $\beta << 1$, the wind of the star will flow along the magnetic lines of force.

1. Heating. The electron temperature in those parts of the mantle in which magnetic lines of force are present is higher than what it would have been if the temperature was that which may be found by considering the flux of radiation coming from the photosphere and enforcing the constraint of radiative equilibrium. This is so because additional energy, believed to originate in the small-scale irregular motions which occur in the high-beta plasma of the photosphere is deposited in these parts of the mantle. The photospheric motions are believed to be driven by the differential rotation of the star. One way that additional heat may be deposited in the plasma of the mantle when magnetic lines of force are present is as a result of the dissipation of the MHD waves which will be created by the buffeting of the foot points of the lines of force by moving elements of plasma in the photosphere; another is by the annihilation of magnetic flux. We shall look at some studies of this problem which have been done for the Sun. The results of these studies may be applicable to hot stars. Magnetic fields which are so small that they cannot be measured in the case of early-type stars using present techniques are adequate to generate significant heating.

Ionson (1982, 1983) has reviewed the theory of the heating of the solar corona, and he has shown how use of the analog of an LRC circuit (i.e., one containing inductance, resistance, and capacitance) enables one to understand the way in which mechanical energy in the high-beta plasma of the photosphere is transferred to the low-beta plasma of a coronal loop. Ionson noted that the magnetic structures in the solar corona can act like "high-Q" circuits, absorbing energy from the noise of mechanical energy which comes from the photosphere. In particular, he showed (as others have done) that the coronal loops act as resonant structures where MHD waves of certain characteristic properties are efficiently dissipated.

Ionson has pointed out the advantages of thinking about the electrodynamic reaction of the low-beta plasma of the mantle to the irregular motions which occur in the high-beta photosphere in terms of: (1) the ability of a magnetoplasma to store electric and kinetic energy (i.e., in terms of its capacitance, *C*), (2) its ability to store magnetic energy (i.e., in terms of its inductance, *L*), and (3) its ability to convert electrodynamic energy into heat (i.e., in terms of its resistance, *R*). Thinking in terms of the solar problem, he specifies the size of the coronal part of a loop by a field-aligned length, l_{cor}, and the size of the chromospheric part by a field-aligned length, l_{chrom}. The field-aligned-length scale of the photospheric velocity fields which interact with the magnetoplasma of the loop are denoted by l_{photo}. As a convenience, it is assumed that the cross-field dimension of the magnetic structure, l, is the same in all three regions. It is postulated that the magnetic field, B_o, which links the plasma in the photosphere to that in the mantle, is generated by a dynamo which is external to the photosphere and to the loop. A representation of what Ionson envisages is shown in Figure 8-2.

Study of the situation shown in Figure 8-2 and use of the analog furnished by an LRC circuit lead to the conclusion that the coronal heating function, the field-aligned current of electrons, and the cross-field plasma flow are explicit functions of the amount of power at the resonant frequency of the loop in the power spectrum of the field of motions in the photosphere. Ionson works in terms of the macroscopic properties of the magnetoplasma, and he finds it unnecessary to specify particular mechanisms for dissipating MHD waves in "high-Q" resonators such as many of the solar loops. In "low-Q" resonators such as the coronal loops of young active regions and the loops of solar bright points, it may be necessary

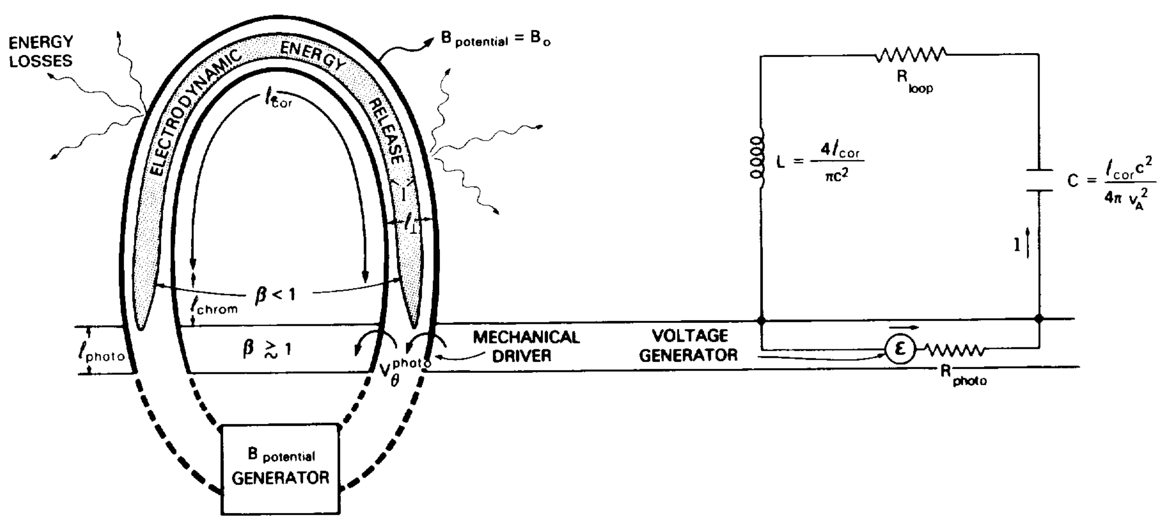


Figure 8-2. The magnetic-loop system and equivalent LRC circuit considered by Ionson (1982).

to consider the details of the dissipating mechanisms if a full understanding of the situation is to be obtained (Ionson, 1984).

It turns out that the heating function for a "high-Q" coronal loop can be written as

$$E_H = 6.23 \times 10^9 \, (T_{eff}/\ell_{cor}^{\,2}) \left\{ R_{loop}/(R_{loop} + R_{photo}) \right\} \langle 0.5 \rho v_\theta^{\,2} \rangle^{photo}_{\omega} \, \mathrm{erg\ cm^{-3}\,s^{-1}} \; . \tag{8-4}$$

Here ω is the resonant frequency of the loop, a quantity which is dependent on the length of the loop, the density distribution in the loop, and the size of the magnetic field. The quantity $\langle 0.5 \ \rho v_\theta^{\,2} \rangle^{photo}_{\omega}$ represents the ensemble average value of the kinetic energy due to azimuthal motions in the photosphere occurring at frequency ω in the power spectrum of the photosphere. (The corresponding circular frequency is $2\pi\omega$.) The ratio, $R_{loop}/(R_{loop} + R_{photo})$, gives the fraction of the power supplied which is absorbed by the loop. This ratio may be evaluated from equations given by Ionson (1982). These equations involve the microscopic properties of the magnetoplasma. The values of β, T, and B enter these expressions as well as the density of electrons and of ions. The heating is the result of irreversible processes which occur when MHD waves dissipate.

From Equation (8-4), we see that long loops will receive less energy per unit volume than small loops if the ensemble average of the kinetic energy in the photosphere at the resonant frequency is not strongly dependent on the frequency in the photospheric power spectrum to which the loop is resonant and if $R_{loop}/(R_{loop} + R_{photo})$ is not sensitively dependent on l_{cor}.

In the case of stars, one may assume, as a first approximation, that the mechanical energy in the photosphere has a white-noise spectrum (i.e., that there is the same power at every frequency of interest). The available estimates of densities and microturbulent velocities in the photospheres of main-sequence and supergiant early-type stars indicate that the available kinetic energy in the photospheres of main-sequence and supergiant stars is about the same for all stars with $T_{eff} > 20000$ K. Consequently, other things being comparable, E_H will tend to increase by about a factor 2 as one goes from type B2 to type O5. A large range in E_H is not expected as one considers early B-type stars, O stars, and Wolf-Rayet stars. Ionson (1982) has stressed that, because a typical loop acts like a "high-Q" LRC circuit, all of the electrodynamic energy entering a loop will be dissipated as heat.

Expressions for the maximum temperature in a coronal loop having particle densities, length scales, and magnetic fields as in the Sun, and for the base pressure of the loop have been derived by Ionson. These expressions may serve as first approximations for hot stars because it seems that the particle densities and magnetic fields in the mantles of hot stars may not differ greatly from those observed for the Sun. On the assumption that the exciting spectrum of mechanical energy is white, one may estimate the maximum temperature in a loop from

$$T_{max} = 2.18 \times 10^4 (T_{eff} \langle \mathrm{K.E.} \rangle^{photo})^{2/7} \exp[0.05 \, \ell_{cor}/\ell_p] \ \mathrm{K}, \tag{8-5}$$

and the pressure in a loop from

$$P_{base} = (7.8 \times 10^3/\ell_{cor})(T_{eff} \langle \mathrm{K.E.} \rangle^{photo})^{6/7} \exp[0.21 \, \ell_{cor}/\ell_p] \ \mathrm{dyn\ cm^{-2}} \; . \tag{8-6}$$

Here $\langle \mathrm{K.E.} \rangle^{photo}$ is the ensemble average kinetic energy per unit area of the motions giving the "white noise" in the photosphere, and l_p is the pressure scale height in the photosphere. Since l_p decreases as log g increases, T_{max} and P_{base} will be larger in main-sequence stars than in supergiants if the other factors are comparable. The range in T_{max} and P_{base} will not be so great for early-type stars as for solar-type stars because the range in photospheric pressure scale height is smaller among the early-type stars than it is for solar-type stars of a given effective temperature.

The dependence of T_{max} in the loop on the effective temperature of the star is relatively small. However, P_{base} increases rather rapidly as T_{eff} increases. In a given star, long loops (l_{cor} large) should tend to have higher T_{max} than do short loops if the exciting spectrum is white. On the other hand, P_{base} should tend to decrease as l_{cor} increases by a modest amount. If the ratio, l_{cor}/ℓ_p, remains more or less constant when going from solar-type stars to early-type stars (as it may well do), the chief differences in T_{max} and P_{base} will be determined by the differences in the product, T_{eff} $\langle K.E.\rangle^{photo}$, between the two types of star. This product may be expected to be larger for early-type stars than it is for solar-type stars, perhaps by a factor of the order of 5.

Ionson's (1982) theory can be used to find numerical factors for Equations (8-5) and (8-6), which correspond to estimated conditions in the mantles of early-type stars, but these factors will not be worked out here because of lack of firm input data. Equations (8-5) and (8-6) lead one to expect maximum temperatures in the loops of early-type stars like those of the Sun and base pressures which are not widely different from the solar case. The base pressures will be lower than the base pressures found for the Sun because the pressure in the photosphere of an early-type star, main-sequence star, or supergiant is less than the pressure in the photosphere of the Sun. The pressure at the base of a loop cannot exceed the pressure in the photosphere.

A study of MHD wave propagation in inhomogeneous plasmas and heating by the mechanism of resonant absorption has been carried out by Rae and Roberts (1982) with a view to understanding better the phenomena seen to occur in the solar corona. The topics of their study are of interest also for understanding the mantles of early-type stars because the plasma densities and magnetic field densities and their distribution in space may be comparable in the mantles of early-type stars to what is observed for the Sun.

Rae and Roberts have noted that propagation effects occur in inhomogeneous plasmas (plasmas in which the sound and Alfvén speeds are functions of position) which are not apparent in unbounded uniform media. They emphasize the importance of the cusp resonance which arises when MHD waves propagate in a compressible plasma, and they suggest that this resonance may provide an important means of heating the solar corona by slow-mode MHD waves.

In a plasma sufficiently uniform that the approximations of geometric optics are valid, it is possible to identify Alfvén, slow, and fast waves as distinct modes. In strongly inhomogeneous plasma, however, the three MHD modes are not necessarily distinct. Rae and Roberts show that, in the inhomogeneous case, the MHD waves may be locally fast or slow, propagating or tunneling. There is no separate Alfvén motion, although there will be particular locations at which a tunneling magnetoacoustic wave propagates along the magnetic field with the local Alfvén speed.

Rae and Roberts note that in coronal loops the cusp resonance can be excited by the action of slow body waves which propagate into it or by slow surface waves which tunnel into it. This mechanism offers possibilities for heating the corona of the Sun in addition to those provided by the Alfvén resonances which are basic to the problem studied by Ionson (1982, 1983).

Another study of a particular mechanism for heating an inhomogeneous corona is that of Heyvaerts and Priest (1983). These authors have considered the coronal heating which may occur as a result of the phase mixing of shear Alfvén waves in areas where the magnetic field is inhomogeneous. Shear Alfvén waves have a displacement perpendicular to the direction of the inhomogeneity in the magnetic field. Heyvaerts and Priest prove that, after a sufficiently long time, phase mixing can ensure the dissipation of all the wave mechanical energy that a loop can pick up from its excitation by white noise, a result which is in agreement with the findings of Ionson and deduced by him in terms of the properties of "high-Q" circuits. Heyvaerts and Priest conclude that phase mixing is the process most able to ensure the dissipation of shear Alfvén waves in loops and in open regions of strong reflectivity.

Unlike magnetoacoustic modes, shear Alfvén waves do not obey a single propagation equation. They propagate independently on each magnetic surface. Because of this independence, the vibrations become more and more out of phase as the propagation proceeds. This has the result that friction grows and energy is dissipated into heat by irreversible processes. The process of energy dissipation by phase mixing is particularly efficient in structures in which multiple reflections set up standing waves (for example, in loops). If the plasma is to pick up energy from the exciting source of mechanical energy, it is necessary that standing shear Alfvén waves be set up. Standing shear Alfvén waves can readily be set up in loops.

Heyvaerts and Priest (1983) also studied the behavior of propagating and standing shear Alfvén waves in the presence of Kelvin-Helmholtz turbulence and tearing mode (magnetic-field-line reconnection) instabilities. They found that propagating shear Alfvén waves appear to be stable against both types of instability, but that standing shear Alfvén waves are unstable to both. Consequently, the parts of a loop in which the magnetic field is inhomogeneous may be in a state of permanent tearing and Kelvin-Helmholtz turbulence. This state of affairs can provide heating, and it is an essential element to consider when evaluating the transport properties of the plasma in the loops and the typical velocities of ions and electrons there.

In the mantles of OB supergiants, some radiatively driven turbulence may be present (see Chapter 7, *Atmospheric Stability: X Rays*). This possibility means that it is important to consider turbulence when evaluating the kinematic viscosity and magnetic diffusity in a magnetic structure of an OB supergiant for use with the theories of resonant heating.

The inhomogeneity of the spatial distribution of the magnetic field is an essential ingredient of the problem studied by Heyvaerts and Priest (1983), as it is for the problems studied by Rae and Roberts (1982) and by Ionson (1982, 1983).

If the theoretical statements of Heyvaerts and Priest are to be valid for a star, conditions must be such that

$$k_{\parallel} a << 1, \tag{8-7}$$

where $k_{\parallel}$ is the wave number of the shear Alfvén waves, and a is the halfwidth of the region in which the inhomogeneity of the magnetic field occurs. From the information given by Heyvaerts and Priest, one sees that

$$k_{\parallel} = \omega / v_A(x) , \tag{8-8}$$

where, in the case of open field lines, ω is the frequency at which the magnetic field lines oscillate as a result of buffeting by turbulent elements of plasma in the photosphere, while in the case in which standing waves occur, ω is the frequency of the standing waves which the magnetic field line can accommodate. The quantity, $v_A(x)$, is the Alfvén velocity at position x. At this position on the surface of the star, the magnetic field has a value $B_o(x)$. Consequently, the Alfvén velocity at x can be found from

$$v_A{}^2(x) = B_o{}^2(x)/4\pi\rho_o(x) . \tag{8-9}$$

Here $\varrho_o(x)$ is the density in the photosphere. Since in the photospheres of early-type stars $\varrho_o \sim 10^{-10}$ g cm^{-3}, one sees that $v_A \sim B_o$ km s^{-1} in the photosphere. Here B_o is expressed in gauss.

If the length of a typical turbulent element in the photosphere is l_{photo} km, then in the case of open regions, ω will be of the order of v_{turb}/l_{photo}. It follows that the theory of Heyvaerts and Priest (1983) can be applied to those coronal-hole-type regions in which the magnetic-field inhomogeneities occur over regions having halfwidths such that

$$a << B_o (l_{photo}/v_{turb}) \text{ km}. \tag{8-10}$$

Typically, in an OB supergiant, v_{turb} may be of the order of 15 km s^{-1}, while in a main-sequence OB star, it may be of the order of 5 km s^{-1}. The size of the turbulent elements is unknown, but they must be small with respect to the stellar radius if the term "microturbulence" is to have meaning. Let us suppose that $l_{photo} \sim 0.01\ R_*$. Since R_* is about 30 $R_\odot$ for an OB supergiant and about 10 $R_\odot$ for a main-sequence OB star (see Part I of this book), we find that we must have

$$a \ll 0.14\ B_o \text{ km} \tag{8-11}$$

if the theory of Heyvaerts and Priest (1983) is to be applied to coronal-hole-type structures on OB stars. The open magnetic structures must have rather sharp edges because it is unlikely that B_o exceeds a few hundred gauss. If it were larger than this, net longitudinal magnetic fields would have been measured for OB stars.

If the theory of Heyvaerts and Priest is to be applied to closed magnetic loops, the halfwidth of the inhomogeneity must be such that

$$a \ll PB_o \text{ km}, \tag{8-12}$$

where P is the period of the standing shear Alfvén waves, and B_o is a typical value for the magnetic field at the inhomogeneity. Possibly P is of the order of 100 seconds and B_o is of the order of 100 gauss. Then the linear size of the inhomogeneities at the edges of coronal loops should be small with respect to 10^4 km. Loops meeting this requirement would have quite sharp edges with respect to the radius of the star, because the radius is about 7×10^6 km in the case of a main-sequence OB star and about 2×10^7 km in the case of a supergiant OB star. If loops are composed of many filaments, the theory of Heyvaerts and Priest can be used for cases in which the cross section of the filaments is of the order of a few times the halfwidths discussed here.

The approaches of Rae and Roberts (1982) and of Heyvaerts and Priest (1983) to modeling the mechanism which causes a plasma to become heated when the magnetic lines of force traversing the plasma are buffeted emphasize what may happen as MHD waves move through the ionized gas. Another way to look at the problem is to focus attention on what may happen to the magnetic lines of force as they are moved around. Studies of this topic for the Sun have been summarized by Parker (1983a, 1983b). Parker shows that, whenever twisted flux tubes are packed closely together, they are subject to dynamical nonequilibrium and to internal neutral-point reconnection. These actions cause a rapid dissipation of the torsion of the flux tubes.

In Parker (1983b), the rate at which work is done on the foot points of tubes of force on the Sun is estimated. Parker finds that the energy transferred to each element of surface containing a magnetic field of B gauss and released as reconnection takes place is

$$P = \frac{w^2}{4\pi}\ \frac{h}{L}\ \frac{B^2}{v} \text{ erg cm}^{-2}\text{ s}^{-1} \tag{8-13}$$

Here w is the typical speed at which the foot points move, h is the width of the flux tube across which reconnection takes place, L is the length of the magnetic line of force, and v is the speed at which reconnection takes place. From consideration of the microscopic properties of the plasma, Parker argues that v should lie in the range

$$Re_{\mathrm{mag}}^{-1/2} \leqslant v/v_A \leqslant (\ln Re_{\mathrm{mag}})^{-1/2} \tag{8-14}$$

where Re_{mag} is the magnetic Reynolds number of the plasma, and v_A is the Alfvén velocity. Parker goes on to argue that rapid reconnection may generate all the energy (about 10^7 erg cm^{-2}) required to provide the observed heating of typical loops in the solar corona. However, the conclusion that rapid reconnection is the dominant mechanism for heating solar loops of the type which Parker specifies is inconsistent with the behavior which may be expected from such loops. This may be seen by considering t_{recon}, the time which it takes for the reconnection of adjacent tubes of magnetic flux.

It is informative to compare the time $t_{recon} = h/v$ to the typical time $t_A = v_A/L$ which is needed for MHD waves to propagate the length of a typical loop. We consider values for the ratio

$$x = t_{recon}/t_A = (h/v)/(L/v_A). \quad (8\text{-}15)$$

Substitution for v from Equation (8-13) and for v_A from Equation (8-9) leads to the relation,

$$x = \frac{P}{w^2 B} (\rho/4\pi)^{-1/2}. \quad (8\text{-}16)$$

Here ϱ is the density (assumed constant) of the plasma in the loop, and P is the energy per unit surface area supplied by the reconnection (see Equation (8-13)).

If reconnection between adjacent flux tubes, each carrying a field B, is to be the dominant heating mechanism and the heating which may take place as the MHD waves propagate through the loop is to be unimportant, then the ratio x should be less than unity (i.e., neighboring tubes of force must reconnect before the MHD waves can move a significant distance along the loop). Demanding that $x << 1$ leads to a contradiction with Parker's conclusion that reconnection is the chief heating mechanism for loops like those which he describes. This can be seen as follows.

The typical plasma properties adopted by Parker are $w = 4 \times 10^4$ cm s^{-1}, $B = 100$ gauss, and $\varrho = 3.32 \times 10^{-15}$g cm^{-3}. Substitution of these values into Equation (8-16) and application of the constraint that x should be smaller than unity shows that the power released by reconnection must fall in the range,

$$P < 2.56 \times 10^3 \text{ erg cm}^{-2}\text{ s}^{-1} \quad . \quad (8\text{-}17)$$

When such a value of P is substituted into Equation (8-13), we find that the velocity of reconnection in the plasma specified by Parker must be greater than 5×10^7 cm s^{-1}.

However, Parker has estimated that v can lie in the range from 25 cm s^{-1} to 1.6×10^7 cm s^{-1} only. We conclude that, if the speed of reconnection lies in its allowed range, given the physical description of a typical loop adopted by Parker, then the power generated by rapid reconnection is a small fraction of the total power which appears to be transferred to the loop as a result of the buffeting of the foot points of the magnetic lines of force by elements of moving plasma in the photosphere. A large fraction of the observed heating in the Sun is best described by mechanisms which take account of the action of MHD waves as they propagate through the loop.

Ionson (1983) has formulated a description of the heating mechanisms for loops which attempts to unify the several approaches which have been taken for representing what occurs. When the loops have such physical properties that they are equivalent to "high-Q" circuits (underdamped systems), the rate of heating is independent of the mechanism by which the heating process is described, whereas when the loops have properties which make them equivalent to "low-Q" LRC circuits (overdamped systems), the rate of heating which occurs is sensitive to the dissipation process which is postulated. It is beyond the scope of this review to go into this subject more deeply. We refer the reader to the literature on the heating of solar coronal loops.

In summary, heating by MHD waves or by the annihilation of magnetic flux involves the transfer of energy to the plasma by irreversible processes. The energy in the organized motion of the plasma, described in terms of MHD waves or that in magnetic flux, disappears, and the random motions of the particles in the plasma are increased. This state of affairs is described by saying that the electron temperature has increased. When waves are dissipated, the momentum associated with the fluid in motion disappears in the sense that it is transferred to the particles of the plasma in such a way that no net component of momentum in any direction is given to the particles. The random motions corresponding to the increased electron temperature may create such a high pressure that the force of gravity cannot contain the heated plasma. A loop may be expected to rise and grow, eventually breaking and releasing the plasma which it contains. This material will escape from the star unless the attractive force of gravity is sufficiently large at the point of release so as to hold the hot plasma under gravitational control. Material which starts out in a coronal-hole-type structure and is heated by the resonant processes or the magnetic flux annihilation processes which take place in such a structure will also escape from the star once the flow velocity exceeds a critical value. (See the discussion of Equation (7-17) and the section on *Characteristic Features of the Stellar Wind Problem.*)

The theory reviewed in Chapter 7 was developed for stars without magnetic fields. This theory points out the principal factors of the wind problem. These factors must be considered whether magnetic fields are present or not. Magnetic fields provide specific mechanisms for the heating and the momentum transfer in the vicinity of the critical point which are necessary for a flow originating in the gravity field of a star to attain supersonic flow velocities and for the material to escape from the star. In order to estimate how far out in the wind acceleration of the flow continues, one must study how momentum in an outward direction is transferred to the particles of the plasma. Information relevant to this problem follows.

2. Momentum Transfer. Here we shall discuss how MHD waves may transfer directed momentum to the plasma particles as a result of the reaction of the plasma of the mantle to the passage of MHD waves. (The effects of radiation pressure are discussed in Chapter 7.) The efficiency with which a specified flux of MHD waves in a spectrum of frequencies drives a wind from a star depends on the character of the microscopic processes which occur as the MHD waves propagate through the parts of the mantle in which magnetic fields are present. Relevant points are how efficiently momentum is transferred from the waves to the particles and how rapidly the waves are attenuated.

Standing MHD waves do not drive winds because, to a first approximation at least, their interaction with the plasma is symmetrical so far as the direction of propagation is concerned. Significant input of directed momentum occurs only in those parts of the plasma in which propagating MHD waves are present.

The results reported here are based on the behavior of propagating MHD waves in plasmas which vary slowly in time and in space. It is assumed that the length and time scales with which the plasma varies are large with respect to the wavelengths of the MHD waves which are present and to the periods of the waves. This is the WKB (Wentzel, Kramers, and Brillouin) approximation. It is now known, however, that heating occurs efficiently in inhomogeneous parts of the magnetoplasma. Therefore, attention should be paid to how the inhomogeneity of the mantle may affect the transfer of momentum from the organized motion of MHD waves to the particles in the mantle of the star.

Waves in a stellar mantle heat the mantle as they are dissipated. They also do work on the expanding gas of a wind, thereby heating the wind and at the same time propelling it outward. Belcher (1971) has worked out the effect of Alfvén waves on the solar wind, and Jacques (1977) has extended this type of treatment to find the stress exerted on the wind by propagating waves of several types. In general, the stress due to waves is anisotropic.

Models for wave-driven stellar winds have been developed for Alfvén waves propagating along radially directed magnetic lines of force. Belcher (1971) and Jacques (1978) have applied the theory to obtain models for the solar wind, while Hartmann and MacGregor (1980) have used the theory to develop some model winds for late-type stars. Additional modeling of the winds of cool stars has been presented by Holzer et al. (1983). These studies show that the rate at which the propagating waves are damped as they move outward is a significant parameter for determining the character of the model wind. Holzer et al. present a self-consistent theory of the frictional damping of Alfvén waves, and they note the importance of *where* the damping occurs for determining the asymptotic properties of the wind.

In the case of Alfvén waves, the component of force (see Equation (7-17)) acting in a radial direction to drive the wind can be expressed as $1/2\ (d\epsilon/dr)$, where ϵ is the wave energy density at the distance r from the center of the star. An explicit formula for $\epsilon(r)$ can be found in the WKB approximation, and this is what is used with appropriate formulations of the conservation equations to make the wind models of Belcher (1971), Jacques (1978), Hartmann and MacGregor (1980), and Holzer et al. (1983).

Hartmann and Cassinelli (1981) and Cassinelli (1982) have published brief reports of attempts to apply the theory of Alfvén-wave-driven winds to obtain an understanding of the winds of Wolf-Rayet stars. The details presented by Cassinelli (1982) are not internally consistent in the framework of wind models developed, for instance, by Jacques and by Hartmann and MacGregor. Furthermore, the problem studied by Cassinelli is unrealistic in that he attempts to provide a source for the force needed to drive a wind which has an unrealistically large density (see Underhill, 1983a). Cassinelli fits his numerical results to what he adopts as the parameters of a Wolf-Rayet star by assigning an ad hoc value to the parameter which describes how rapidly the propagating Alfvén waves dissipate as they move outward. At the time of writing, no internally consistent applications of the theory of wave-driven winds exist which are satisfactory for explaining the observations of the winds of O and Wolf-Rayet stars.

Five different factors have been studied for providing the driving force for the winds of stars. Most model winds made by considering these forces are one-dimensional; in only one study (Kopriva and Jokipii, 1983) is the dependence of the force on the angular coordinates considered. The five factors are:

1. The gas pressure exerted by a hot plasma (Parker, 1958, 1963; Leer and Holzer, 1980; Hammer, 1982a, 1982b; Hearn, 1983; Kopriva and Jokipii, 1983);
2. The force exerted by radiation absorbed in resonance lines (Lucy and Solomon, 1970; Castor, Abbott, and Klein, 1975; Abbott, 1980, 1982; Weber, 1981);
3. The force exerted by waves (Belcher, 1971; Jacques, 1977, 1978; Hartmann and MacGregor, 1980; Holzer et al., 1983; Davila, 1985);
4. Rotation in the presence of a magnetic field (Weber and Davis, 1967; Limber, 1974; Saito, 1974; Nerney, 1980; Barker and Marlborough, 1982; Barker, 1982; Friend and MacGregor, 1984);
5. The gradient in a magnetic field as the magnetic lines of force diverge around a magnetic bubble formed by reconnection (Schluter, 1957; Parker, 1957; Mullan and Ahmad, 1982; Pneuman, 1983).

In some of the investigations which study the effects of rotation, the solutions for the flow of gas from the star which are obtained are unrealistic in the sense that the velocity law in the radial direction is not constrained to go through the critical point.

The properties of a three-dimensional wind flowing from a rotating star and driven by gas pressure have been studied by Kopriva and Jokipii (1983). These authors evade having to consider an energy equation by adopting a polytrope relation between density and pressure, a step which is taken in several of the studies listed above. Rotation of the star causes material to move toward the pole of the star when it is first released. This poleward component of motion is, however, only temporary. Eventually, the material moves outward in essentially a purely radial direction, but in a different angular direction from which it started.

Some properties of a three-dimensional wind from a rotating star have been studied by Underhill and Fahey (1984). The results of Kopriva and Jokipii (1983) confirm the typical three-dimensional paths and motions of ejected parcels of gas found by Underhill and Fahey. The question studied by Underhill and Fahey is: Where does a parcel of gas released from a spot on a rotating star go if, after its release, it experiences such a force in a radial direction that the component of velocity in the radial direction follows a law like that believed to be typical for hot stars (see Equation (7-53)). Rotation of a star distorts the density pattern in a wind from that typical of uniform spherical outflow from a stationary star. Ejection from a spot on or above the photosphere of a rotating star will lead to material lying in the cone of sight and far from the star if the colatitude of the point of ejection is equal to the inclination of the pole of the star.

Studies aimed at providing a model for the solar wind have considered all of the factors listed above except radiation pressure. However, as Pneuman (1983) points out, none of the models are entirely satisfactory. The solar wind is observed to be a mixture of fast and slow streams, as well as of magnetic bubbles. The changing inhomogeneous structure of the solar wind observed from spacecraft clearly indicates the importance of magnetic fields for providing some of the driving force for the solar wind. One result of this is that the solar wind becomes quite chaotic, on a small scale, at distances of more than 25 AU from the Sun, although it becomes homogeneous on the large scale (Tyler et al., 1981; Burlaga, 1983). The streams do not maintain their identity forever.

All of the above factors may influence the winds from hot stars. The findings about the solar wind suggest that it is unlikely to be true that the winds from hot stars are as simple as has been portrayed by models which use radiation pressure as the only driving force, or by those which attempt to model the effects of rotation and magnetic fields in the equatorial plane. The fact that there is a considerable spread in v_∞ among stars of a given T_{eff} suggests that forces in addition to the force of radiation pressure are driving the winds of hot stars (Underhill, 1983c), while the existence of shortward displaced components in the resonance lines suggests that some gas is ejected from particular spots on or above the photosphere (Underhill and Fahey, 1984). The terminal velocity which is reached is a result of the sum of the radial force acting and the action of the torques generated by the rotation of the star.

3. Mass Flux from a Star. The mass flux from a star at any moment in the life of a star is determined by the rate at which gas is released from gravitational control and from confinement by magnetic lines of force. The observed rates of mass loss are averages over the surface of the star of the local rates of mass efflux.

The statement that the plasma is released from a star means that the plasma, whether moving as individual particles or as a parcel of plasma, has achieved such components of motion that it can move to a great distance from the star and eventually meld with the interstellar medium. The conditions which are necessary for this to occur have been discussed in Chapter 7 and above. What will be discussed here are the circumstances which favor the release of plasma from a star. We are particularly interested in the case of hot stars.

Observations of the Sun have shown that the wind from a star which has a magnetic field distributed on the disk of the star in bipolar magnetic regions (regions which act as the foot points for coronal loops) and in unipolar magnetic regions (regions in which coronal holes develop) has three components. They are: (1) streams of rapidly moving material, (2) streams of slowly moving material, and (3) moving parcels of plasma which have been described as magnetic bubbles, coronal transients, or coronal bullets (see, for instance, Karpen et al., 1982), depending on where they are observed in the solar wind and what their density and degree of ionization is.

A coronal hole is a part of the solar corona in which the magnetic lines of force are open, being radially directed. The temperature of the plasma is lower than it is in a coronal loop, although high. Solar coronal holes, like solar coronal loops, form, develop, move about, and dissipate in intervals of time of the order of several weeks to several months. Their behavior appears to be a result of the changes which occur in the surface solar magnetic field (see, for instance, Bohlin and Sheeley, 1978). The character of the magnetic-field topology is the dominant factor which distinguishes coronal holes from coronal loops. The difference in heating which is generated in open and closed magnetic structures by MHD waves will account in a qualitative way for the difference in electron temperature between the plasma in holes and that confined in loops.

Plasma may break free from the confining forces of gravity and of curved magnetic lines of force by receiving an impulse which gives it an outward component of velocity. The material may then experience a continuing acceleration as a result of the action of the five forces listed above. The initial impulse may originate in the "snapping" of magnetic lines of force. That is, curved lines of force may reconnect to form shorter loops, thereby leaving some plasma free from the confining force due to relatively short curved magnetic-field lines.

Observations of the solar mantle have shown that rather frequently some material suddenly begins to move rapidly outward from levels above the chromosphere. It appears, however, that only some of the material in the solar wind is released with the sudden acquisition of a large component of velocity. Most of the material lying at the base of the corona in coronal holes moves outward rather slowly (see, for instance, Orrall et al., 1983). Nevertheless, a significant outward component of motion is detected at levels so deep in the corona that the five driving forces listed above can hardly have had much time in which to work. We can say that outflow may start at those levels in the mantle where the material becomes free from the confining forces of curved magnetic-field lines and only experiences the attractive force of gravity and the expelling forces listed above. A wind is composed only of those particles and parcels which find trajectories satisfying the conservation equations and the boundary conditions which are appropriate for melding with the interstellar medium.

Pneuman (1983) has discussed the release of material to form the high-speed streams which emerge from solar coronal holes, and he suggests that the release points may be identified with very small X-ray bright spots. Pneuman indicates that the reconnection of magnetic-field lines, whether it occurs only on a small linear scale as in the release of high-speed streams or on a large linear scale as may be the case for the release of magnetic bubbles, is the dominant mechanism for generating the solar wind. Magnetic reconnection is a result of the twisting and moving of magnetic lines of forces caused by the buffeting of the foot points of the lines of force by turbulent elements of gas in the photosphere.

Parker (1983a, 1983b) has discussed magnetic reconnection. We have seen in the section on *Heating* that, because magnetic reconnection may take place relatively slowly, magnetic reconnection may not be an important process for heating the average solar loop. However, it may be important for releasing material. Here we suggest that magnetic reconnection across narrow threads of plasma occurs in the mantles of all hot stars and that it is the mechanism which regulates the mass flux from a star. Reconnection causes small bodies of plasma—the "plasmoids" of Pneuman (1983)—to start to move outward between the radially directed lines of force emerging from a unipolar magnetic region. The motion and trajectory of the "plasmoid" once it has been released depends on the sizes of the particular forces active in the stellar mantle at the place where the "plasmoid" is. Some remarks on this aspect of the problem can be found in Underhill and Fahey (1984).

How much of the wind from a hot star is similar to the streams of plasma emerging from solar coronal holes, how much is similar to the streams of plasma emerging from the part of the solar surface covered by coronal loops, and how much is in the form of discrete parcels of plasma ejected from a few spots above the surface of the star will depend on the character of the surface of the star (i.e., on how much of the surface is covered by unipolar magnetic regions, how much is covered by bipolar magnetic regions, and how much is essentially free from emerging magnetic flux).

Underhill and Fahey (1984) have given reasons for believing that bipolar magnetic regions occur on the surfaces of hot stars, particularly OB supergiants, while Underhill (1983a) has noted that the surfaces of Wolf-Rayet stars seem to be predominantly covered by magnetic loop structures.

Very possibly, the disks of Be stars may be formed as magnetic loops on the surface of a rapidly rotating B star twist, tangle, and mat together owing to the torques generated by the rotation of the star. The physical conditions (density, electron temperature, and degree of ionization) in the disk would be the result of the heating and cooling processes which occur as the magnetic lines of force tangle and MHD waves, generated by turbulence in the photosphere at the foot points of the lines of force, propagate. The temporal changes of the spectra of Be stars can be attributed to the growth and decay of a disk as the amount of magnetic flux emerging from the photosphere of a rotating B star and the distribution of this flux on the surface of the star change. Why cool disks should be generated around some rotating B stars and not around others is a question which cannot be answered until further information is available about how magnetic fields are continuously (or discontinuously) generated in and expelled from stars of different masses and ages, and how MHD waves propagate along tangled lines of force.

All of the above statements are qualitative. The development of numerical models confirming or rejecting the pictures sketched above should be a future goal for work aimed at understanding the spectra of hot stars and determining the physical properties of hot stars.

9

SUMMARY OF PROCESSES INFLUENCING THE SPECTRA OF O AND WOLF-RAYET STARS

In the preceding chapters, we have seen that the strengths of the emission lines and of almost all of the strong absorption lines in the spectra of O and Wolf-Rayet stars, the emission of X rays and an infrared and radio excess, and the outflow of a wind cannot be interpreted in a self-consistent way if one uses only classical stellar-atmosphere theory based on the constraints of hydrostatic, radiative, and statistical equilibrium. In Chapter 7, we have reviewed the existing theory of the atmospheres of hot stars, noting its successes and its failures, while in Chapter 8, we have explored in depth how expanding the group of assumptions on which the theory of stellar spectra is based to include the presence of relatively small inhomogeneous magnetic fields assists us greatly in understanding what is observed.

Radiatively driven turbulence may be the cause of some of the spectroscopic phenomena which are observed, but it seems most likely that the prime causes of the high temperatures and outflow deduced to be present in the mantles of O, Of, and Wolf-Rayet stars are the heating and momentum transfer which result from turbulent motions in the photosphere buffeting the foot points of magnetic lines of force which thread the photosphere. We postulate that the needed magnetic lines of force originate in dynamos situated in the He^+ ionization layer which occurs in the deepest parts of the photospheres of O and Wolf-Rayet stars. The primary source of the motion which is described as turbulence and which is invoked to generate magnetohydrodynamic (MHD) waves (which then propagate through the mantles of the O, Of, and Wolf-Rayet stars) is the rotation of the star. The viscosity of the rotating fluid which forms the boundary layer of the star is believed to generate a turbulent helical flow in the photosphere, a flow which is parallel to the surface of the star. This flow then creates dynamos in the He^+ convection zone and generates MHD waves.

Radiatively driven turbulence (see Chapter 7) may be generated when radiation traverses a low-density ionized gas in which there is a velocity gradient. However, it remains to be shown that the sound waves needed in the theories of radiatively driven turbulence can propagate from the photosphere, or deeper layers in which they may be generated, to the mantle of the star. The necessary velocity gradient for the theories of radiatively driven turbulence can be generated by the action of MHD waves (see Chapter 8). Once a significant velocity gradient has been generated, the effects of radiation pressure in lines may augment the velocity gradient generated by the action of MHD waves.

Five questions were posed at the end of Chapter 6. We shall now present brief answers to these questions:

1. The ultimate source for the nonradiative energy and momentum which are released and which cause a mantle to form is the rotation of the star.

2. The energy and momentum which originate in the rotation of the star are transferred to the gas in the mantle of the star by the actions of MHD waves as they propagate along the magnetic-field lines which traverse the mantle. The magnetic-field lines are generated in dynamos which arise as helical motion generated by the rotation of the star occurs in the weak general magnetic field of the star. It is likely that the magnetic-field lines pass through the photosphere only in narrow regions concentrated at the edges of turbulent eddies, as is observed to be the case in the Sun.

3. The release of nonradiative energy and momentum by the actions of MHD waves is not significant in the photosphere because, in the parts of the photosphere where magnetic-field lines are present, the plasma beta (ratio of gas pressure to magnetic pressure) is large. In the parts of the photosphere free from magnetic-field lines, no MHD waves are present.

4. Nonradiative phenomena in stellar atmospheres are important for understanding the evolution of stars in each case that the empirically selected criteria for spectral type are spectroscopic features dependent for their strength and appearance on the presence of nonradiatively generated heat and momentum in the mantle. It is commonly supposed that a spectral type may always be used as an unambiguous indicator of the primary parameters of the theory of stellar evolution. We recall that the primary parameters of the theory of stellar evolution are the mass, radius, effective temperature, and composition of the star. It has been shown in Chapter 7 that these basic parameters cannot be deduced unambiguously from the spectra of O, Of, and Wolf-Rayet stars when one uses solely classical theory. The observed appearance of O, Of, and Wolf-Rayet spectra depends sensitively on the amount of nonradiative energy and momentum in the mantle of the star, as well as on the size of the mantle and the state of motion there. Consequently, it is very important to take note of the occurrence of nonradiative phenomena in the atmospheres of hot stars if one is to understand the evolution of the hot stars correctly. Information about the masses, radii, and effective temperatures of O and Wolf-Rayet stars may be found in Part I of this book. The theoretical studies reviewed in Part III indicate that Population I O, Of, and Wolf-Rayet stars have normal solar composition. In some subluminous O stars, the relative abundance of helium to hydrogen is not normal. It appears that gravitational settling of helium may have taken place in some cases, while in others, most of the hydrogen-rich envelope may have been lost.

5. The observation of nonradiative phenomena in the atmosphere of O, Of, and Wolf-Rayet stars does not give guidance concerning the stage of evolution of isolated single O, Of, and Wolf-Rayet stars. Spectra which are classified as O, Of, and Wolf-Rayet are found among the normal massive young stars of Population I and among the subluminous stars. The latter stars can rightfully be regarded as old highly

evolved objects. The information presented in Part III of this book indicates that the spectroscopic phenomena which typify O, Of, and Wolf-Rayet spectra are most likely due to the heating and momentum transfer in a mantle caused by the rotation of a hot star which has a weak magnetic field in its boundary layer. The boundary-layer magnetic field may be primordial in origin, or it may have been generated during the evolution of the star. Until the theory of stellar evolution reaches such a state of completeness that one can predict the occurrence of magnetic flux in the boundary layer of a hot star, more specifically in the He^+ ionization zone of its photosphere, one cannot use only the observation of spectroscopic phenomena which suggest heating and momentum transfer primarily generated by the actions of MHD waves in the mantle of the star to deduce surely the stage of evolution of a hot star. Other independent information such as the composition of the atmosphere, the luminosity of the star, its effective temperature, and its mass are needed before one can estimate the stage of evolution of the star.

In summary, the appearance of the spectra of O, Of, and Wolf-Rayet stars is strongly conditioned by the actions of nonradiative heating and nonradiative momentum transfer in the mantle of the star. Because this is so, an O, Of, or Wolf-Rayet spectral type is not an unambiguous guide to the stage of evolution of the star.

It may be possible to deduce interesting evolutionary information from spectral types of classes O, Of, and Wolf-Rayet when the theory of stellar evolution has been extended to include a prediction of when in their lifetimes massive model stars (and those of low mass) having high effective temperatures may have magnetic flux in their subsurface He^+ ionization zones, and when they may be rotating at such a speed that helical flow motion is generated in the boundary layer of the star. The latter flow will create subsurface dynamos in the general magnetic field with a consequent expulsion of magnetic lines of force from the star and the formation of a mantle having the properties which are deduced for the mantles of O, Of, and Wolf-Rayet stars. In Chapter 8, it is explained how the expulsion of magnetic lines of force can account for the mantles of O, Of, and Wolf-Rayet stars, including the rate of mass loss from each star.

REFERENCES

Abbott, D.C. 1978, *Astrophys. J.*, **225**, 893.

Abbott, D.C. 1980, *Astrophys. J.*, **242**, 1183.

Abbott, D.C. 1982, *Astrophys. J.*, **259**, 282.

Abbott, D.C., Bieging, J.H., and Churchwell, E. 1981, *Astrophys. J.*, **250**, 645.

Abbott, D.C., Bieging, J. H., and Churchwell, E. 1982, in *Wolf-Rayet Stars: Observations, Physics, Evolution*, ed. C.W.H. de Loore and A.J. Willis (Dordrecht: Reidel), p. 215.

Abbott, D.C., Bieging, J.H., Churchwell, E., and Cassinelli, J.P. 1980, *Astrophys. J.*, **238**, 196.

Abney, W. de W. 1877, *Mon. Not. Roy. Astr. Soc.*, **37**, 278.

Allen, C.W. 1973, *Astrophysical Quantities* (London: Athlone Press).

Andriesse, C.D. 1978, *Astrophys. Space Sci.*, **61**, 205.

Andriesse, C.D. 1980a, *Astrophys. Space Sci.*, **67**, 461.

Andriesse, C.D. 1980b, *Astrophys. Space Sci.*, **72**, 167.

Andriesse, C.D. 1981, in *Effects of Mass Loss on Stellar Evolution*, ed. C. Chiosi and R. Stalio (Dordrecht: Reidel), p. 213.

Auer, L.H., and Heasley, J.N. 1976, *Astrophys. J.*, **205**, 165.

Auer, L.H., and Mihalas, D. 1969, *Astrophys. J.*, **158**, 641.

Auer, L.H., and Mihalas, D. 1972, *Astrophys. J. Supplement*, **24**, 193.

Auer, L.H., and van Blerkom, D. 1972, *Astrophys. J.*, **178**, 175.

Avrett, E.H., and Krook, M. 1963a, *Astrophys. J.*, **137**, 874.

Avrett, E.H., and Krook, M. 1963b, *J. Quant. Spectrosc. Rad. Transfer*, **3**, 107.

Barker, P.K. 1982, in *Be Stars*, ed. M. Jaschek and H.G. Groth (Dordrecht: Reidel), p. 485.

Barker, P.K., Landstreet, J.D., Marlborough, J.M., Thompson, I., and Maza, J. 1981, *Astrophys. J.*, **250**, 300.

Barker, P.K., and Marlborough, J.M. 1982, *Astrophys. J.*, **254**, 297.

Barlow, M.J. 1979, in *Mass Loss and Evolution of O-Type Stars*, ed. P. S. Conti and C.W.H. de Loore (Dordrecht: Reidel), p. 119.

Barlow, M.J., and Cohen, M. 1977, *Astrophys. J.*, **213**, 737.

Barlow, M.J., Smith, L.J., and Willis, A.J. 1981, *Mon. Not. Roy. Astr. Soc.*, **196**, 101.

Baschek, B., Kudritzki, R.P., and Scholz, M. 1980, in *Proc. Second European IUE Conference*, ed. B. Battrick and J. Mort, ESA SP-157, p. 315.

Baschek, B., and Scholz, M. 1971, *Astron. Astrophys.*, **15**, 285.

Beals, C.S. 1929, *Mon. Not. Roy. Astr. Soc.*, **90**, 202.

Belcher, J.W. 1971, *Astrophys. J.*, **168**, 509.

Bernacca, P.L., and Bianchi, L. 1979, *Astron. Astrophys.*, **75**, 61.

Bertout, C. 1977, *Astron. Astrophys.*, **58**, 153.

Bhatia, A.K., and Underhill, A.B. 1986, *Astrophys. J. Supplement*, **60**, 323.

Bohlin, J.D., and Sheeley, N.R. 1978, *Solar Phys.*, **56**, 125.

Böhm, K.-H. 1973, in *Extended Atmospheres and Circumstellar Matter in Spectroscopic Binary Systems*, ed. A.H. Batten (Dordrecht: Reidel), p. 148.

Bowen, I.S. 1947, *Pub. Astron. Soc. Pacific*, **59**, 196.

Burlaga, L.F. 1983, *J. Geophys. Res.*, **88**, 6085.

Burlaga, L.F., Schwenn, R., and Rosenbauer, H. 1983, *Geophys. Res. Letters*, **10**, 413.

Canfield, R.C. 1969, *Astrophys. J.*, **157**, 425.

Cannon, C.J., and Thomas, R.N. 1977, *Astrophys. J.*, **211**, 910.

Cardona, O. 1978, Ph.D. Thesis, Univ. Colorado, Boulder, Colorado.

Carlberg, R.G. 1980, *Astrophys. J.*, **241**, 1131.

Cassinelli, J.P. 1971, *Astrophys. J.*, **165**, 265.

Cassinelli, J.P. 1979, *Ann. Rev. Astron. Astrophys.*, **17**, 275.

Cassinelli, J.P. 1982, in *Wolf-Rayet Stars: Observations, Physics, Evolution*, ed. C.W.H. de Loore and A.J. Willis (Dordrecht: Reidel), p. 173.

Cassinelli, J.P., and Hartmann, L. 1975, *Astrophys. J.*, **202**, 718.

Cassinelli, J.P., and Hartmann, L. 1977, *Astrophys. J.*, **212**, 418.

Cassinelli, J.P., and Olson, G.L. 1979, *Astrophys. J.*, **229**, 304.

Cassinelli, J.P., Olson, G.L., and Stalio, R. 1978, *Astrophys. J.*, **220**, 573.

Cassinelli, J.P., Waldron, W.L., Sanders, W.T., Harnden, F.R., Rosner, R., and Vaiana, G.S. 1981, *Astrophys. J.*, **250**, 677.

Castor, J.I. 1970, *Mon. Not. Roy. Astr. Soc.*, **149**, 111.

Castor, J.I. 1974a, *Astrophys. J.*, **189**, 273.

Castor, J.I. 1974b, *Mon. Not. Roy. Astr. Soc.*, **169**, 279.

Castor, J.I., Abbott, D.C., and Klein, R.I. 1975, *Astrophys. J.*, **195**, 157.

Castor, J.I., Abbott, D.C., and Klein, R.I. 1976, in *Physique des Mouvements dans les Atmosphères Stellaires*, ed. R. Cayrel and M. Steinberg (Paris: Editions du CNRS), p. 363.

Castor, J.I., and Cassinelli, J.P. 1973, *Astrophys. J.*, **179**, 189.

Castor, J.I., and Lamers, H.J.G.L.M. 1979, *Astrophys. J. Supplement*, **39**, 481.

Castor, J.I., Lutz, J.H., and Seaton, M.J. 1982, *Mon. Not. Roy. Astr. Soc.*, **194**, 547.

Castor, J.I., and Nussbaumer, H. 1972, *Mon. Not. Roy. Astr. Soc.*, **155**, 293.

Castor, J.I., and Simon, T. 1983, *Astrophys. J.*, **265**, 304.

Castor, J.I., and van Blerkom, D. 1970, *Astrophys. J.*, **161**, 485.

Cayrel, R. 1963, *Compt. Rend.*, **257**, 3309.

Chandrasekhar, S. 1939, *An Introduction to the Study of Stellar Structure* (Chicago: Univ. Chicago Press).

Chandrasekhar, S. 1950 *Radiative Transfer* (Oxford: Clarendon Press), p. 65.

Chandrasekhar, S. 1961, *Hydrodynamic and Hydromagnetic Stability* (Oxford: Clarendon Press).

Cherepashchuk, A.M. 1975, *Soviet Astr.-A.J.*, **19**, 47.

Cohen, M., Barlow, M.J., and Kuhi, L.V. 1975, *Astron. Astrophys.*, **40**, 291.

Collard, H.R., Mihalov, J.D., and Wolfe, J.H. 1982, *J. Geophys. Res.*, **87**, 2203.

Conti, P.S., and Frost, S.A. 1977, *Astrophys. J.*, **212**, 728.

Conti, P.S., and Garmany, C.D. 1980, *Astrophys. J.*, **238**, 190.

Conti, P.S., and Leep, E.M. 1974, *Astrophys. J.*, **193**, 113.

Conti, P.S., Leep, E.M., and Perry, D.N. 1983, *Astrophys. J.*, **268**, 228.

Cowling, T.G. 1981, *Ann. Rev. Astron. Astrophys.*, **19**, 115.

Davila, 1985, *Astrophys J.*, **291**, 328.

de Groot, M., and Underhill, A.B. 1964, *Bull. Astron. Inst. Netherlands*, **17**, 280.

Dufton, P.L. 1972, *Mon. Not. Roy. Astr. Soc.*, **159**, 79.

Dumont, S., and Heidmann, N. 1973, *Astron. Astrophys.*, **27**, 273.

Ebbets, D. 1979, *Astrophys. J.*, **227**, 510.

Feautrier, P. 1967, *Ann. Astrophys.*, **30**, 125 and 604.

Feautrier, P. 1968, *Ann. Astrophys.*, **31**, 257.

Felli, M., and Panagia, N. 1981, in *Effects of Mass Loss on Stellar Evolution*, ed. C. Chiosi and R. Stalio (Dordrecht: Reidel), p. 179.

Friend, D.B., and MacGregor, K.B. 1984, *Astrophys. J.*, **282**, 591.

Garcia, J.D., and Mack, J.E. 1965, *J. Opt. Soc. Am.*, **55**, 654.

Garmany, C.D., Olson, G.L., Conti, P.S., and Van Steenberg, M. 1981, *Astrophys. J.*, **250**, 660.

Gathier, R., Lamers, H.J.G.L.M., and Snow, T.P. 1981, *Astrophys. J.*, **247**, 173.

Gerasimoric, B.P. 1934, *Mon. Not. Roy. Astron. Soc.*, **94**, 737.

Glasco, H.P., and Zirin, H. 1964, *Astrophys. J. Supplement*, **9**, 193.

Grachev, S.I., and Grinen, V.P. 1975, *Astrophys.*, **11**, 20.

Griem, H.R. 1963, *Phys. Rev.*, **131**, 1170.

Groot, M., and Thé, P.S. 1983, *Astron. Astrophys.*, **120**, 89.

Gruschinske, J., Hunger, K., Kudritzki, R.P., and Simon, K.P. 1980, in *Proc. Second European IUE Conference*, ed. B. Battrick and J. Mort, ESA SP-157, p. 311.

Gruschinske, J., and Kudritzki, R.P. 1979, *Astron. Astrophys.*, **77**, 341.

Hackwell, J.A., Gehrz, R.D., and Smith, J.R. 1974, *Astrophys. J.*, **192**, 383.

Hamann, W.-R. 1980, *Astron. Astrophys.*, **84**, 342.

Hamann, W.-R. 1981, *Astron. Astrophys.*, **93**, 353.

Hammer, R. 1982a, *Astrophys. J.*, **259**, 767.

Hammer, R. 1982b, *Astrophys. J.*, **259**, 779.

Hammer, R. 1984, *Astrophys. J.*, **280**, 780.

Hammer, R., Endler, F., and Ulmschneider, P. 1983, *Astron. Astrophys.*, **120**, 141.

Harnden, F.R., Branduardi, G., Elvis, M., Gorenstein, P., Grindlay, J., Pye, J.P., Rosner, R., Topka, K., and Vaiana, G.S. 1979, *Astrosphys. J. (Letters)*, **234**, L51.

Hartmann, L. 1978a, *Astrophys. J.*, **221**, 193.

Hartmann, L. 1978b, *Astrophys. J.*, **224**, 520.

Hartmann, L., and Cassinelli, J.P. 1977, *Astrophys. J.*, **215**, 155.

Hartmann, L., and Cassinelli, J.P. 1981, *Bull. Amer. Astron. Soc.*, **13**, 785.

Hartmann, L., and MacGregor, K.B. 1980, *Astrophys. J.*, **242**, 260.

Hearn, A.G. 1975a, *Astron. Astrophys.*, **40**, 277.

Hearn, A.G. 1975b, *Astron. Astrophys.*, **40**, 355.

Hearn, A.G. 1979, in *Mass Loss and Evolution of O-Type Stars*, ed. P.S. Conti and C.W.H. de Loore (Dordrecht: Reidel), p. 169.

Hearn, A.G. 1981, in *Effects of Mass Loss on Stellar Evolution*, ed. C. Chiosi and R. Stalio (Dordrecht: Reidel), p. 125.

Hearn, A.G. 1983, *Astron. Astrophys.*, **116**, 296.

Hearn, A.G., and Vardavas, I.M. 1981a, *Astron. Astrophys.*, **98**, 230.

Hearn, A.G., and Vardavas, I.M. 1981b, *Astron. Astrophys.*, **98**, 246.

Heyvaerts, J., and Priest, E.R. 1983, *Astron. Astrophys.*, **117**, 220.

Hogg, D.E. 1982, in *Wolf-Rayet Stars: Observations, Physics, Evolution*, ed. C.W.H. de Loore and A.J. Willis (Dordrecht: Reidel), p. 221.

Hollweg, J.V. 1981, in *The Sun as a Star*, ed. S.D. Jordan, NASA SP-450, Chapter 15.

Holzer, T.E., Flå, T., and Leer, E. 1983, *Astrophys. J.*, **275**, 808.

Holzer, T., and Leer, E. 1980, *J. Geophys. Res.*, **85**, 4665.

Huang, S.-S., and Struve, O. 1960, in *Stellar Atmospheres*, ed. J.L. Greenstein (Chicago: Univ. Chicago Press), p. 321.

Hummer, D.G. 1976, in *Be and Shell Stars*, ed. A. Slettebak (Dordrecht: Reidel), p. 281.

Hundt, E., Kodaira, K., Schmid-Burgk, J., and Scholz, M. 1975, *Astron. Astrophys.*, **41**, 37.

Hunger, K., Gruschinske, J., Kudritzki, R.P., and Simon, K.P. 1981, *Astron. Astrophys.*, **95**, 244.

Hutchings, J.B. 1976, *Astrophys. J.*, **203**, 438.

Hutchings, J.B., and von Rudloff, I.R. 1980, *Astrophys. J.*, **238**, 909.

Inglis, D.R., and Teller, E. 1939, *Astrophys. J.*, **90**, 439.

Ionson, J.A. 1982, *Astrophys. J.*, **254**, 318.

Ionson, J.A. 1983, *Astrophys. J.*, **271**, 778.

Ionson, J.A. 1984, *Astrophys. J.*, **276**, 357.

Jacques, S.A. 1977, *Astrophys. J.*, **215**, 942.

Jacques, S.A. 1978, *Astrophys. J.*, **226**, 632.

Jordan, S.D. 1981, in *The Sun as a Star*, ed. S.D. Jordan, NASA SP-450, Chapter 12.

Kahn, F.D. 1981, *Mon. Not. Roy. Astr. Soc.*, **196**, 641.

Kalkofen, W. 1970, in *Spectrum Formation in Stars with Steady-State Extended Atmospheres*, ed. H.G. Groth and P. Wellmann, NBS Special Pub. 332, p. 120.

Kamp, L.W. 1976, *Statistical Equilibrium Calculations for Silicon in Early-Type Model Atmospheres*, NASA Tech. Report TR R-455.

Kamp, L.W. 1978, *Astrophys. J. Supplement*, **36**, 143.

Karpen, J.T., Oran, E.S., Mariska, J.T., Boris, J.P., and Breuckner, G.E. 1982, *Astrophys. J.*, **261**, 375.

Klein, R.I., and Castor, J.I. 1978, *Astrophys. J.*, **220**, 902.

Kopp, R.A. 1981, in *The Sun as a Star*, ed. S.D. Jordan, NASA SP-450, Chapter 16.

Kopriva, D.A., and Jokipii, J.R. 1983, *Astrophys. J.*, **267**, 62.

Kosirev, N.A. 1934, *Mon. Not. Roy. Astr. Soc.*, **94**, 430.

Kron, G.E., and Gordon, K. 1950, *Astrophys. J.*, **111**, 454.

Kuan, P., and Kuhi, L.V. 1975, *Astrophys. J.*, **199**, 148.

Kudritzki, R.P. 1976, *Astron. Astrophys.*, **52**, 11.

Kudritzki, R.P., Hunger, K., Gruschinske, J., and Simon, K.P. 1980, in *Proc. Second European IUE Conference*, ed. B. Battrick and J. Mort, ESA SP-157, p. 307.

Kudritzki, R.P., and Simon, K.P. 1978, *Astron. Astrophys.*, **70**, 653.

Kunasz, P.B. 1980, *Astrophys. J.*, **237**, 819.

Kunasz, P.B., Hummer, D.G., and Mihalas, D. 1975, *Astrophys. J.*, **202**, 92.

Kuperus, M., Ionson, J.A., and Spicer, D.S. 1981, *Ann. Rev. Astron. Astrophys.*, **19**, 7.

Kurucz, R.L. 1979, *Astrophys. J. Supplement*, **40**, 1.

Kurucz, R.L., Peytremann, E., and Avrett, E.H. 1974, *Blanketed Model Atmospheres for Early-Type Stars*, Smithsonian Institution, U.S. Govt. Printing Office, Washington, D.C. 20402.

Kurucz, R.L., and Schild, R.E. 1976, in *Be and Shell Stars*, ed. A. Slettebak (Dordrecht: Reidel), p. 377.

Lamers, H.J.G.L.M. 1981, in *Effects of Mass Loss on Stellar Evolution*, ed. C. Chiosi and R. Stalio (Dordrecht: Reidel), p. 181.

Lamers, H.J.G.L.M., and Morton, D.C. 1976, *Astrophys. J. Supplement*, **32**, 715.

Lamers, H.J.G.L.M., and Rogerson, J.B. 1978, *Astron. Astrophys.*, **66**, 417.

Ledoux, P. 1965, in *Stellar Structure*, ed. L.H. Aller and D.B. McLaughlin (Chicago: Univ. Chicago Press), p. 499.

Ledoux, P., and Walraven, Th. 1958, in *Hb. d. Phys.*, **51**, ed. S. Flugge (Berlin: Springer-Verlag), p. 353.

Leer, E., and Holzer, T. 1980, *J. Geophys. Res.*, **85**, 4681.

Leibacher, J., and Stein, R.F. 1976, in *Physique des Mouvements dans les Atmosphères* Stellaires, ed. R. Cayrel and M. Steinberg (Paris, Editions du CNRS), p. 69.

Limber, D.N. 1974, *Astrophys. J.*, **192**, 429.

Lucy, L.B. 1971, *Astrophys. J.*, **163**, 95.

Lucy, L.B. 1982, *Astrophys. J.*, **255**, 286.

Lucy, L.B., and Solomon, P.M. 1970, *Astrophys. J.*, **159**, 879.

Lucy, L.B., and White, R.L. 1980, *Astrophys. J.*, **241**, 300.

Lyong, L.V. 1967, *Soviet Astr.-A.J.*, **11**, 224.

MacGregor, K.B., Hartmann, L., and Raymond, J.C. 1979, *Astrophys. J.*, **231**, 514.

Marlborough, J.M., and Roy, J.R. 1970, *Astrophys. J.*, **160**, 221.

Martens, P.C.H. 1979, *Astron. Astrophys.*, **75**, L7.

Marti, F., and Noerdlinger, P.N. 1977, *Astrophys. J.*, **215**, 247.

Martin, W.C. 1960, *J. Res. Nat. Bur. Stds., A*, **64**, 19.

McCluskey, G.E., and Kondo, Y. 1981, *Astrophys. J.*, **246**, 464.

McCluskey, G.E., Kondo, Y., and Morton, D.C. 1975, *Astrophys. J.*, **201**, 607.

McDonald, F.B., Schardt, A.W., and Trainor, J.H. 1980, *J. Geophys. Res.*, **85**, 5813.

McDonald, F.B., Trainor, J.H., Mihalov, J.D., Wolfe, J.H., and Weber, W.R. 1981, *Astrophys. J. (Letters)*, **246**, L165.

Méndez, R.H., Kudritzki, R.P., Gruschinske, J., and Simon, K.P. 1981, *Astron. Astrophys.*, **101**, 323.

Menzel, D.H. 1929, *Pub. Astron. Soc. Pacific*, **41**, 344.

Menzel, D.H., Aller, L. H., and Baker, J. G. 1938, *Astrophys. J.*, **88**, 313.

Michaud, G. 1977, *Highlights of Astronomy*, **4**, Part II, 177.

Mihalas, D. 1965, *Astrophys. J. Supplement*, **9**, 321.

Mihalas, D. 1969, *Astrophys. J. (Letters)*, **156**, L155.

Mihalas, D. 1970, *Stellar Atmospheres* (San Francisco: W.H. Freeman and Company).

Mihalas, D. 1972a, *Astrophys. J.*, **176**, 139.

Mihalas, D. 1972b, *Non-LTE Model Atmospheres for B and O Stars*, NCAR-TN/STR-76.

Mihalas, D. 1972c, *Astrophys. J.*, **177**, 115.

Mihalas, D. 1978, *Stellar Atmospheres*, 2nd Edition (San Francisco: W.H. Freeman and Company).

Mihalas, D., Barnard, A.J., Cooper, J., and Smith, E.W. 1974, *Astrophys. J.*, **190**, 315.

Mihalas, D., and Conti, P.S. 1980, *Astrophys. J.*, **235**, 515.

Mihalas, D., Frost, S.A., and Lockwood, G.W. 1975, *Pub. Astron. Soc. Pacific*, **87**, 153.

Mihalas, D., Heasley, J.N., and Auer, L.H. 1975, *A Non-LTE Model Stellar Atmosphere Computer Program*, NCAR-TN/STR-104.

Mihalas, D., and Hummer, D.G. 1973, *Astrophys. J.*, **179**, 827.

Mihalas, D. and Hummer, D.G. 1974, *Astrophys. J. Supplement*, **28**, 343.

Mihalas, D., and Kunasz, P.B. 1978, *Astrophys. J.*, **219**, 635.

Milne, E.A. 1926, *Mon. Not. Roy. Astron. Soc.*, **86**, 459.

Moffat, A.F.J., and Michaud, G. 1981, *Astrophys. J.*, **251**, 133.

Moore, C.E. 1965, NSRDS-NBS3, Section 1.

Moore, C.E. 1972, NSRDS-NBS3, Section 6.

Morton, D.C. 1967a, *Astrophys. J.*, **147**, 1017.

Morton, D.C. 1967b, *Astrophys. J.*, **150**, 535.

Morton, D.C., Jenkins, E.B., and Bohlin, R.C. 1968, *Astrophys. J.*, **154**, 661.

Morton, D.C., and Wright, A.E. 1978, *Mon. Not. Roy. Astr. Soc.*, **182**, 47P.

Morton, D.C., and Wright, A.E. 1979, in *Mass Loss and Evolution of O Stars*, ed. P.S. Conti and C.W.H. de Loore (Dordrecht: Reidel), p. 155.

Mullan, D.J. 1984, *Astrophys. J.*, **283**, 303.

Mullan, D.J. and Ahmad, I.A. 1982, *Solar Phys.*, **75**, 347.

Münch, G. 1950, *Astrophys. J.*, **112**, 266.

Nelson, G.D., and Hearn, A.G. 1978, *Astron. Astrophys.*, **65**, 223.

Nerney, S. 1980, *Astrophys. J.*, **242**, 723.

Noerdlinger, P.D., and Rybicki, G.B. 1974, *Astrophys. J.*, **193**, 651.

Nugis, T. 1975, in *Variable Stars and Stellar Evolution*, ed. V.E. Sherwood and L. Plaut (Dordrecht: Reidel), p. 291.

Nussbaumer, H. 1982, *Astron. Astrophys.*, **110**, L1.

Nussbaumer, H., Schmutz, W., Smith, L.J., and Willis, A.J. 1982, *Astron. Astrophys. Supplement*, **47**, 257.

Oegerle, W.R., and van Blerkom, D. 1976a, *Astrophys. J.*, **206**, 150.

Oegerle, W.R., and van Blerkom, D. 1976b, *Astrophys. J.*, **208**, 453.

Oke, J.B. 1954, *Astrophys. J.*, **120**, 22.

C-6

Olnon, F.M. 1975, *Astron. Astrophys.*, **39**, 217.

Olson, G.L. 1978, *Astrophys. J.*, **226**, 124.

Olson, G.L. 1981, *Astrophys. J.*, **245**, 1054.

Olson, G.L., and Castor, J.I. 1981, *Astrophys. J.*, **244**, 179.

Olson, G.L., and Ebbets, D. 1981, *Astrophys. J.*, **248**, 1021.

Orrall, F.Q., Rottman, G.J., and Klimchuk, J.A. 1983, *Astrophys. J.*, **266**, L65.

Pallavicini, R., Golub, L., Rosner, R., Vaiana, G.S., Ayres, T., and Linsky, J.L. 1981, *Astrophys. J.*, **248**, 279.

Panagia, N., and Felli, M. 1975, *Astron. Astrophys.*, **39**, 1.

Parker, E.N. 1957, *Astrophys. J. Supplement*, **3**, 51.

Parker, E.N. 1958, *Astrophys. J.*, **128**, 664.

Parker, E.N. 1963, *Interplanetary Dynamical Processes* (New York: Interscience Pub.).

Parker, E.N. 1981, *Astrophys. J.*, **251**, 266.

Parker, E.N. 1983a, *Astrophys. J.*, **264**, 635.

Parker, E.N. 1983b, *Astrophys. J.*, **264**, 642.

Pecker, J.C. 1950, *Ann. Astrophys.*, **13**, 433.

Pecker, J.C. 1965, *Ann. Rev. Astron. Astrophys.*, **3**, 135.

Pecker, J.C., Praderie, F., and Thomas, R.N. 1973, *Astron. Astrophys.*, **29**, 289.

Perinotto, M., and Panagia, N. 1981, in *Effects of Mass Loss on Stellar Evolution*, ed. C. Chiosi and R. Stalio (Dordrecht: Reidel), p. 41.

Peterson, D.M., and Scholz, M. 1971, *Astrophys. J.*, **163**, 51.

Pneuman, G.W. 1983, *Astrophys. J.*, **265**, 468.

Rae, I.C., and Roberts, B. 1982, *Mon. Not. Roy. Astr. Soc.*, **201**, 1171.

Rosner, R., Golub, L., Coppi, B., and Vaiana, G.S. 1978a, *Astrophys. J.*, **222**, 317.

Rosner, R., Tucker, W.H., and Vaiana, G.S. 1978b, *Astrophys. J.*, **220**, 643.

Rottman, G.J., Orrall, F.Q., and Klimchuk, J.A. 1982, *Astrophys. J.*, **260**, 326.

Rudkjøbing, M. 1947, *Pub. Mind. Medd. Københavns Obs.*, No. 145.

Rublev, S.V. 1961, *Soviet Astr.-A.J.*, **4**, 780.

Rublev, S.V. 1963, *Soviet Astr.-A.J.*, **6**, 686.

Rublev, S.V. 1964a, *Soviet Astr.-A.J.*, **7**, 492.

Rublev, S.V. 1964b, *Soviet Astr.-A.J.*, **8**, 172.

Rublev, S.V. 1972a, *Isv. Spets. Astrofiz. Obs.*, **4**, 3.

Rublev, S.V. 1972b, *Isv. Spets. Astrofiz. Obs.*, **4**, 18.

Rublev, S.V. 1972c, *Isv. Spets. Astrofiz. Obs.*, **4**, 32.

Rublev, S.V. 1975, in *Variable Stars and Stellar Evolution*, ed. V.E. Sherwood and L. Plaut (Dordrecht: Reidel), p. 259.

Rumpl, W.M. 1980, *Astrophys. J.*, **241**, 1055.

Rybicki, G.B. 1970, in *Spectrum Formation in Stars with Steady-State Extended Atmospheres*, ed. H.G. Groth and P. Wellmann, NBS Special Pub. 332, p. 87.

Rybicki, G.B., and Hummer, D.G. 1978, *Astrophys. J.*, **219**, 654.

Sahade, J. 1980, *The Wolf-Rayet Stars* (Paris: College de France), p. 67.

Saito, M. 1974, *Pub. Astron. Soc. Japan*, **26**, 103.

Sanders, W.T., Cassinelli, J.P., and van der Hucht, K.A. 1982, in *Wolf-Rayet Stars: Observations, Physics, Evolution*, ed. C.W.H. de Loore and A.J. Willis (Dordrecht: Reidel), p. 589.

Schluter, A. 1957, in *Radio Astronomy*, ed. H.C. van de Hulst (Cambridge: Cambridge Univ. Press), p. 356.

Schmid-Burgk, J., and Scholz, M. 1975, *Astron. Astrophys.*, **41**, 41.

Seaquist, E.R. 1976, *Astrophys. J. (Letters)*, **203**, L35.

Seward, F.D., Forman, W.R., Giacconi, R., Griffiths, R.E., Harnden, F.R., Jones, C., and Pye, J.P. 1979, *Astrophys. J. (Letters)*, **234**, L55.

Simon, K.P. 1979, Thesis, Kiel University.

Simon, K.P., Gruschinske, J., Hunger, K., and Kudritzki, K.P. 1980, in *Proc. Second European IUE Conference*, ed. B. Battrick and J. Mort, ESA SP-157, p. 305.

Smith, L.F. 1973a, in *Wolf-Rayet and High-Temperature Stars*, ed. M.K.V. Bappu and J. Sahade (Dordrecht: Reidel), p. 15.

Smith, L.F. 1973b, in *Wolf-Rayet and High-Temperature Stars*, ed. M.K.V. Bappu and J. Sahade (Dordrecht: Reidel), p. 228.

Smith, L.J., and Willis, A.J. 1982, *Mon. Not. Roy. Astr. Soc.*, **201**, 451.

Smith, L.J., and Willis, A.J. 1983, *Astron. Astrophys. Supplement*, **54**, 229.

Sobolev, V.V. 1947, *Moving Envelopes of Stars*, Leningrad State University (in Russian); 1960 transl. by S. Gaposchkin (Cambridge: Harvard Univ. Press).

Sobolev, V.V. 1958, in *Theoretical Astrophysics*, ed. V.A. Ambartzumian (New York: Pergamon Press), Chapter 8.

Sparks, W.M., and Fischel, D. 1971, *Partition Functions and Equations of State in Plasmas*, NASA SP-3066.

Spiegel, E.A. 1976, in *Physique des Mouvements dans les Atmosphères Stellaires*, ed. R. Cayrel and M. Steinberg (Paris: Editions du CNRS), p. 19.

Stein, R.F. 1981, *Astrophys. J.*, **246**, 966.

Stoeckley, T.R., and Mihalas, D. 1973, *Limb-Darkening and Rotational Broadening of Neutral Helium and Ionized Magnesium Line Profiles in Early-Type Stars*, NCAR-TN/STR-84.

Strömgren, B. 1940, *Pub. Mind. Medd. Københavns Obs.*, No. 127.

Strömgren, B. 1944, *Pub. Mind. Medd. Københavns Obs.*, No. 144.

Struve, O., and Elvey, C.T. 1934, *Astrophys. J.*, **79**, 409.

Surdej, J. 1979, *Astron. Astrophys.*, **73**, 1.

Takada, M. 1977, *Pub. Astron. Soc. Japan*, **29**, 439.

Tanzi, E.G., Tarenghi, M., and Panagia, N. 1981, in *Effects of Mass Loss on Stellar Evolution*, ed. C. Chiosi and R. Stalio (Dordrecht: Reidel), p. 51.

Thomas, R.N. 1949, *Astrophys. J.*, **109**, 500.

Thomas, R.N. 1973, *Astron. Astrophys.*, **29**, 297.

Thomas, R.N. 1977, *Highlights of Astronomy*, **4**, Part II, p. 143.

Thomas, R.N. 1979, in *Mass Loss and Evolution of O-Type Stars*, ed. P.S. Conti and C.W.H. de Loore (Dordrecht: Reidel), p. 215.

Traving, G. 1957, *Zs. f. Astrophys.*, **41**, 215.

Tyler, G.L., Vesecky, M.A., Plume, H.T.H., and Barnes, A. 1981, *Astrophys. J.*, **249**, 318.

Uesugi, A. and Fukuda, I. 1970, *Contrib. Inst. Astrophys. Univ. Kyoto and Kwasan Obs.*, No. 189.

Underhill, A.B. 1949, *Mon. Not. Roy. Astr. Soc.*, **109**, 562.

Underhill, A.B. 1950, *Pub. Mind. Medd. Københavns Obs.*, No. 151.

Underhill, A.B. 1951, *Pub. Dom. Astrophy. Obs.*, **8**, 357.

Underhill, A.B. 1955, *J. Roy. Astr. Soc. Canada*, **49**, 27; *Contrib. Dom. Astrophys. Obs.*, No. 43.

Underhill, A.B. 1959, *Pub. Dom. Astrophys. Obs.*, **11**, No. 8, 209.

Underhill, A.B. 1962, *Pub. Dom. Astrophys. Obs.*, **11**, No. 23.

Underhill, A.B. 1966, *The Early-Type Stars*, (Dordrecht: Reidel).

Underhill, A.B. 1968a, *Bull. Astron. Inst. Netherlands*, **19**, 500.

Underhill, A.B. 1968b, *Ann. Rev. Astron. Astrophys.*, **6**, 39.

Underhill, A.B. 1968c, in *Wolf-Rayet Stars*, ed. K.B. Gebbie and R.N. Thomas, NBS Special Pub. 307, p. 183.

Underhill, A.B. 1972, *Vistas in Astronomy*, **13**, 169.

Underhill, A.B. 1980a, *Astrophys. J.*, **239**, 220.

Underhill, A.B. 1980b, *Astrophys. J. (Letters)*, **240**, L153.

Underhill, A.B. 1981, *Astrophys. J.*, **244**, 963.

Underhill, A.B. 1982, *Astrophys. J.*, **263**, 741.

Underhill, A.B. 1983a, *Astrophys. J.*, **265**, 933.

Underhill, A.B. 1983b, *Astrophys. J.*, **266**, 718.

Underhill, A.B. 1983c, *Astrophys. J. (Letters)*, **268**, L127.

Underhill, A.B. 1984a, *Astrophys. J.*, **276**, 583.

Underhill, A.B. 1984b, *Astrophys. J.*, **287**, 874.

Underhill, A.B. 1986, *Pub. Astron. Soc. Pacific*, **98**, 897.

Underhill, A.B., and de Groot, M. 1964, *Bull. Astron. Inst. Netherlands*, **17**, 453.

Underhill, A.B., and de Groot, M. 1965, *Rech. Astron. Obs. Utrecht*, **17**, No. 3.

Underhill, A.B., Divan, L., Prévot-Burnichon, M.-L., and Doazan, V. 1979, *Mon. Not. Roy. Astr. Soc.*, **189**, 601; microfiche *MN* **189**/1.

Underhill, A.B., and Doazan, V. 1982, *B. Stars With and Without Emission Lines*, NASA SP-456.

Underhill, A.B., and Fahey, R.P. 1984, *Astrophys. J.*, **280**, 712.

Unsöld, A. 1942, *Zs. f. Astrophys.*, **21**, 22.

Unsöld, A. 1944, *Zs. f. Astrophys.*, **23**, 100.

Unsöld, A. 1955, *Physik der Sternatmosphären*, (Berlin: Springer-Verlag).

Vaiana, G.S., Cassinelli, J.P., Fabbiano, G., Giacconi, R., Golub, L., Gorenstein, P., Haisch, B.M., Harnden, F.R., Johnson, H.M., Linsky, J.L., Maxson, C.W., Mewe, R., Rosner, R., Seward, F., Topka, K., and Zwaan, C. 1981, *Astrophys. J.*, **245**, 163.

van Blerkom, D. 1973, in *Wolf-Rayet and High-Temperature Stars*, ed. M.K.V. Bappu and J. Sahade (Dordrecht: Reidel), p. 165.

van Blerkom, and Patton, G.Y. 1972, *Astrophys. J.*, **177**, 547.

van der Hucht, K.A., Conti, P.S., Lundström, I., and Stenholm, B. 1981, *Space Sci. Rev.*, **28**, 1.

Vardavas, I.M., and Hearn, A.G. 1981, *Astron. Astrophys.*, **98**, 241.

Vauclair, G., and Vauclair, S. 1977, *Highlights of Astronomy*, **4**, Part II, p. 193.

Vogt, R.E., Chenette, D.L., Cummings, A.C., Garrard, T.L., Stone, E.C., Schardt, A.W., Trainor, J.H., Lal, N., and McDonald, F.B. 1981, *Science*, **212**, 231.

Walborn, N.R. 1971, *Astrophys. J.*, **23**, 257.

Waldron, W.L. 1984, *Astrophys. J.*, **282**, 256.

Weber, E.J., and Davis, L.D. 1967, *Astrophys. J.*, **148**, 217.

Weber, S.V. 1981, *Astrophys. J.*, **243**, 954.

Wentzel, D.G. 1981, in *The Sun as a Star*, ed. S.D. Jordon, NASA SP-450, Chapter 14.

Wesemael, F., Auer, L.H., Van Hoorn, H.M., and Savedoff, M.P. 1980, *Astrophys. J. Supplement*, **43**, 159.

Willis, A.J., and Wilson, R. 1978, *Mon. Not. Roy. Astr. Soc.*, **182**, 559.

Wolff, S.C. 1983, *The A-Type Stars: Problems and Perspectives*, NASA SP-463.

Wolff, S.C., Edwards, S., and Preston, G.W. 1982, *Astrophys. J.*, **252**, 322.

Wright, A.E., and Barlow, M.J. 1975, *Mon. Not. Roy. Astr. Soc.*, **170**, 41.

SUBJECT INDEX

LIST OF CONTRIBUTING AUTHORS

Dietrich Baade	*European Southern Observatory* *Karl-Schwarzschild Strasse 2* *D-8046 Garching* *Federal Republic of Germany*
Peter S. Conti	*Joint Institute for Laboratory Astrophysics* *University of Colorado* *Boulder, CO 80309*
Lucienne Divan	*Institut d'Astrophysique* *98 bis, Boulevard Arago* *75014 Paris, France*
Catharine D. Garmany	*Joint Institute for Laboratory Astrophysics* *University of Colorado* *Boulder, CO 80309*
Huib F. Henrichs	*Astronomical Institute* *University of Amsterdam* *Roetersstraat 15* *1018 WB Amsterdam* *The Netherlands*
Rolf Kudritzki	*Institut für Astronomie und Astrophysik* *der Universität München* *8000 München 80* *Federal Republic of Germany*
Adrian Pauldrach	*Institut für Astronomie und Astrophysik* *der Universität München* *8000 München 80* *Federal Republic of Germany*

Marie-Louise Prévot-Burnichon — *Observatoire de Marseilles*
2 Place de Verrier
F-13248 Marseilles CDX 04
France

Joachim Puls — *Institut für Astronomie und Astrophysik*
der Universität München
8000 München 80
Federal Republic of Germany

Anne B. Underhill — *Department of Geophysics and Astronomy*
University of British Columbia
Vancouver, B.C. VGT 1W5
Canada