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CONTINUED ULTRAVIOLET STUDIES OF SOME Be STARS OF LATER TYPE

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ULTRAVIOLET SPECTRA OF SOME BRIGHT LATER-TYPE Be STARS AND A-F SHELL STARS

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ABSTRACT

Anomalous ionization (C IV and Si IV) is seen in IUE spectra of Be stars as late as B8, and occurs also in standard stars of similar spectral type. Asymmetrical lines suggesting mass loss are present in all the Be stars and several of the standard stars as well, with no obvious correlation with v sin i. Emission shoulders are present in the Mg II lines of two B5e stars but not in Be stars of later type. Again there is no correlation with v sin i. The A-F shell stars show rich Fe II absorption spectra in the ultraviolet, in one case with velocity structure.

I. INTRODUCTION

Early ultraviolet spectroscopic observations of Be stars (e.g. Marlborough and Snow 1976) concentrated on the stars of early type, demonstrating mass-loss effects in their spectra. More recently, these investigations have been extended to cooler Be stars. Thus, Snow (1981) found mass-loss effects in the spectra of Be stars as late as B6, and Marlborough and Peters (1982) detected anomalous ionization (C IV and Si IV) in Be stars as late as B8. This investigation concentrates on the ultraviolet spectra of the later-type Be stars and A-F shell stars, with emphasis on mass-loss effects and anomalous ionization.

II. THE PROGRAM

High-dispersion IUE spectra of 13 bright later-type Be stars and A-F shell stars were obtained in both the short and long wavelength regions, plus 8 standard stars of corresponding spectral types. These are listed in Tables 1a and 1b, with spectral types and rotational velocities from a recent study of the brighter Be stars (Slettebak 1982). The program and standard stars were selected to include both sharp-lined and broad-lined stars of various spectral types between B3 and F0.

Special attention was paid to resonance lines, for possible mass-loss effects, and lines which indicated anomalous ionization in earlier investigations. These included Si II 1190.418, 1193.284, 1260.418, 1264.730; Si III 1206.510; C II 1334.532, 1335.708; Si IV 1393.755, 1402.769; C IV 1548.202, 1550.774; Al II 1670.787; Al III 1854.716, 1862.790; Fe II multiplets 1, 62, and 63; and Mg II 2795.523, 2802.698. Some of these are illustrated in Figures 1-4.

III. RESULTS AND DISCUSSION

A. The Be Stars of Later Type.

The approach taken in this investigation is that of spectral classification: standard non-emission-line stars with MK spectral types close to those of the program stars were observed along with the program stars for direct comparisoh purposes. Furthermore, both sharp-lined and broad-lined standard stars were included, to match as closely as possible the sharp-lined (interpreted as pole-on, according to the rotational model) Be stars and broad-lined (absorption-line shell-type) Be stars. An unexpected result is that the standard star ultraviolet spectra are almost as interesting as the Be spectra, from the point of view of mass loss and anomalous ionization.

1. Anomalous Ionization

According to Marlborough and Peters (1982), the coolest spectral type at which C IV or Si IV would be expected under radiative equilibrium conditions, either in the photosphere or in a cool circumstellar envelope, is B0 and B2, respectively. They observed both ions in IUE spectra of Be stars as late as B8, however. My results, obtained with different stars, confirm their conclusion: the B8e and B7e stars β Cyg B and η Tau have both C IV (see Fig. 1) and Si IV lines in their spectra. On the other hand, C IV seems not to be present in the spectrum of the B5e star β Psc nor in the B6.5e star ϕ And (see Fig. 1), though both are pole-on stars like η Tau. Si IV (see Fig. 2) is present in all the B5-6e stars as well as the later-type stars shown in Fig. 1.

Among the standard non-emission-line stars, anomalous ionization effects also exist. Thus, C IV is seen weakly in the B8 standard ι And (Fig. 1) but is very weak or not present in the B5 standards ρ Aur and ψ^2 Aqr. Si IV, on the hand, is visible in the spectra of all three standard stars.

2. Mass-Loss

Fig. 1 shows that the C IV doublet is asymmetrical and shifted toward shorter wavelengths in both η Tau and β Cyg B. The same is true for the shell stars 48 Lib, ψ Per, and o And. Fig. 2 shows that the Si IV doublet is present and asymmetrical for the B5-6e stars shown there. The lines are also asymmetrical for the B6.5-8e stars shown in Fig. 1. The Al III resonance doublet is also strongly asymmetrical in the spectrum of ψ Per. Thus, mass-loss effects appear to be present in the spectra of both shell and pole-on Be stars of spectral types B3 to B8.

Interestingly, the standard non-emission line stars also show asymmetrical lines, suggestive of mass-loss effects. Thus, Fig. 2 shows the Si IV lines in the sharp-lined B5 standard ρ Aur to be strongly asymmetrical, as are the Al III resonance lines in the broad-lined B5 standard ψ^2 Aqr. In Fig. 3, the Mg II resonance lines are clearly asymmetrical in the spectrum of ρ Aur.

3. Emission Lines

Emission wings in the Mg II and Fe II resonance lines were observed in nearly all the Be stars of earlier type studied with IUE by Dachs (1980). Fig. 3 shows the resonance doublet of Mg II in the spectrum of a number of Be and standard stars. Note that the lines have emission shoulders in the spectra of both β Psc and ψ Per. Both are of spectral type B5, but β Psc is a pole-on star and ψ Per a shell star. The only other emission lines detected in this investigation are the short-wavelength shoulders of the Fe I multiplet 1 lines in ψ Per, but this is not a certain identification.

B. The A-F Shell Stars

The four A-F shell stars in Table 1a are identified as shell stars in the optical wavelength region because of the simultaneous presence of rotationally-broadened lines, from the underlying star, and sharp absorption lines from ground states or low-lying metastable levels, presumably from some kind of shell. All four show rich Fe II absorption spectra in the ultraviolet (especially from multiplets 1, 62, and 63) which are considerably stronger than in the standard star of corresponding type. Fig. 4 shows a portion of the spectrum of one of the A-type shell stars, β Pic, with the A5 standard δ Cas. Note that the Fe II lines are not only stronger and deeper in the shell star, but asymmetrical, indicating velocity structure.

It is a pleasure to acknowledge the very kind hospitality shown to me during a stay at the Institute for Astronomy of the University of Vienna, where much of this work was done. I am grateful also to Ted Snow for very helpful discussions, Ken Carpenter for much help with computer aspects of this investigation, Bob Wing for permission to use his spectra of β Cyg B, Charlie Wu and Al Holm for their spectra of η U Ma, and the IUE team at Goddard for their help and cooperation.

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TABLE 1a. PROGRAM STARS

Name	Sp. Type	v sin i (km s^{-1})
48 Lib	B3:IV:e	400
ψ Per	B5 IIIe	280
β Psc	B5 Ve	100
ο And	B6 IIIe	260
φ And	B6.5 IIIe	80
η Tau	B7 IIIe	140
β Cyg B	B8 Ve	250
28 Tau	B8 (V:)e	320
1 Del	B8-9e	280
β Pic	A5 IV-shell	120
21 Vul	A5 IV-shell	200
φ Leo	A7 IV-shell	230
14 Com	F0 III-shell	200

TABLE 1b. STANDARD STARS

Name	Sp. Type	v sin i (km s^{-1})
η U Ma	B3 V	150
ρ Aur	B5 V	90
ψ^2 Aqr	B5 V	280
α Leo	B7 V	280
ι And	B8 V	90
α Lyr	A0 V	< 10
θ Leo	A2 V	15
δ Cas	A5 III-IV	100

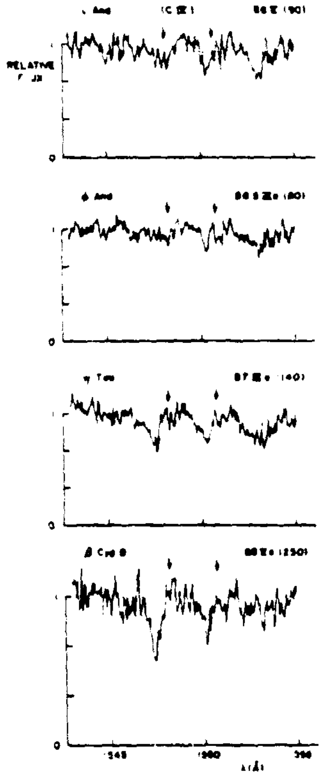


Fig. 1. The C IV resonance doublet in the spectra of B6.5-8 stars.

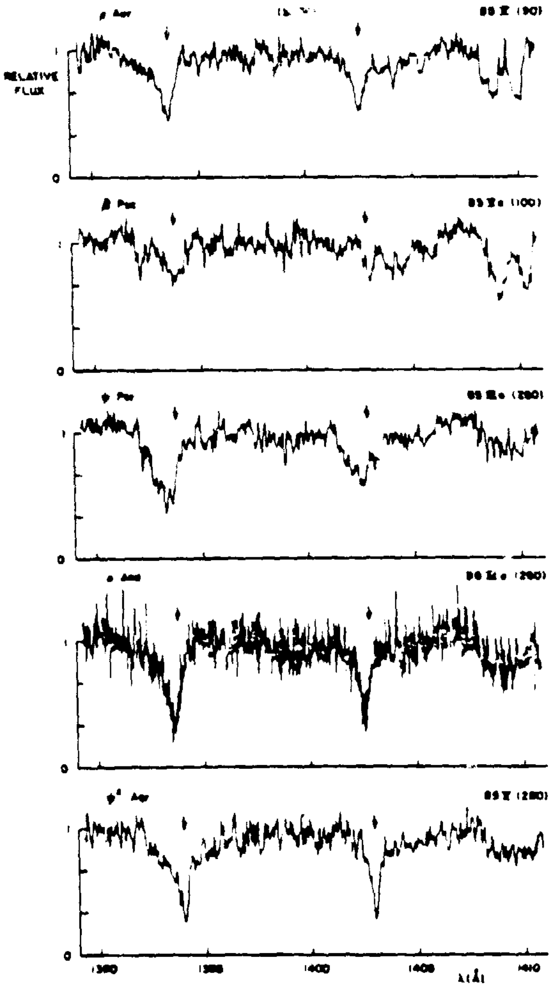
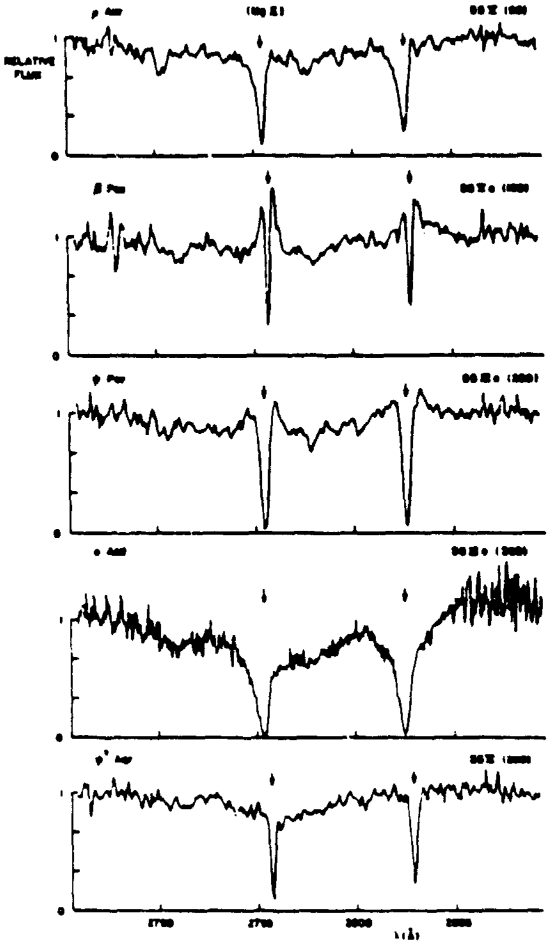


Fig. 2. The Si IV resonance doublet in the spectra of B5-6 stars.



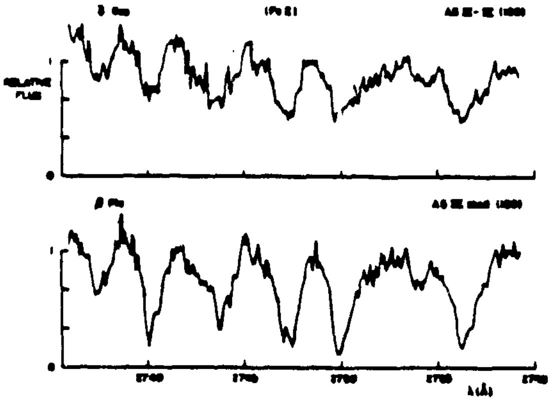


Fig. 4. Fe II multiplet 62 and 63 lines in the spectra of an A5 standard and an A5 shell star.

Fig. 3. (left). The Mg II resonance doublet in the spectra of B5-6 stars.

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ULTRAVIOLET SPECTROSCOPIC OBSERVATIONS OF SOME Be STARS OF LATER TYPE AND A–F TYPE SHELL STARS

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ABSTRACT

High-dispersion *IUE* spectra of 18 later type Be and A–F type shell stars plus eight standard non–emission-line stars were analyzed for anomalous ionization and mass-loss effects. We find that superionization exists in Be stars of the latest spectral subtypes (B8–B9) and is strongest in the Be shell stars, but does not seem to be present in the A–F type shell stars. We also observe superionization in normal B-type stars as late as B5 and possibly B7–B8, but not later. Asymmetrical or violet-displaced resonance lines, suggesting mass loss, were observed in all of our Be program stars but one, and also in a number of our standard stars, but not in the A–F type shell stars. Lower limits to the mass-loss rates computed from Si IV lines range between 5.3×10^{-12} and 3.5×10^{-11} $M_\odot$ yr^{-1}, with the Be shell stars showing the largest values. Terminal velocities measured from the asymmetrical or violet-displaced resonance-line profiles show no clear correlation with ionization potential. Mass loss is correlated with luminosity and effective temperature in the sense that the B-type stars show mass-loss effects while the A-type stars do not, but seems to be uncorrelated with rotation. Emission was observed only in the Mg II resonance doublet for two B5e stars in our sample. All of the A–F type shell stars show strong Fe II and Mg II absorption spectra relative to standard stars of similar spectral type. We conclude from our results that (1) although Be stars are nearly indistinguishable from normal B stars in the ultraviolet, there are differences in degree; and (2) the hot circumstellar envelope surrounding Be stars which gives rise to superionized lines and mass loss appears not to be as nearly equatorially confined as the cool circumstellar envelope.

Subject headings: stars: Be — stars: circumstellar shells — stars: mass loss — ultraviolet: spectra

I. INTRODUCTION

Early ultraviolet spectroscopic observations of Be stars concentrated on the stars of early type (see, e.g., papers in *IAU Symposium 70* [1976]). More recently, these observations have been extended to cooler Be stars. Thus, Snow (1981) found mass-loss effects in the ultraviolet spectra of Be stars as late as B6, while Marlborough and Peters (1982) and Doazan (1982) detected anomalously high ionization (superionization) in Be stars as late as B7–B8. This investigation concentrates on the Be stars of later type, with observations also of the A- and F-type shell stars, which are believed to represent an extension of the Be phenomenon to later types. The object is to ascertain the degree to which mass loss and superionization effects persist to later types in the Be and A–F type shell stars, to compare the ultraviolet spectra of the program stars with those of normal, nonemission standard stars of similar spectral types, and to search for correlations of the effects found with stellar parameters such as $v \sin i$, luminosity, and effective temperature.

II. OBSERVATIONS

High-dispersion *IUE* spectra of 16 program stars and seven standard non–emission-line stars were obtained in both the short (~1200–1900 Å) and long (~1900–3200 Å) wavelength regions during several US 2 shifts in 1980–1982. Archival *IUE* spectra for two additional program stars and one additional standard star were obtained from the National Space Science Data Center (NSSDC).

The 18 program stars are listed in Table 1A in order of spectral type, with spectral types and rotational velocities ($v \sin i$) from the listing by Slettebak (1982). Stars with narrow lines (small $v \sin i$) were included, as well as stars with broad lines (large $v \sin i$). The former may be thought of as "pole-on" stars, if Be stars are interpreted according to Struve's (1931) rotational model, while the latter are then viewed nearly equatorially. These broad-lined stars also often show "shell spectra" (sharp absorption lines which arise from ground states or meta-

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[1] Guest Investigator, *International Ultraviolet Explorer* Satellite.

TABLE 1A

THE PROGRAM STARS

HD	Name/HR	Spectral Type	$v \sin i$ (km s^{-1})	SWP	LWR	Date of Observation
142983	48 Lib	B3: IV:e-shell	400	11212	9829	1981 Feb 1
22192	ψ Per	B5 IIIe-shell	280	10388	9071	1980 Oct 17
217891	β Psc	B5 Ve	100	10386	9069	1980 Oct 17
217675	o And	B6 IIIe-shell	260	15508	11983	1981 Nov 15
183656[a]	HR 7415	B6e-shell	300	14478	11064	1981 Jul 15
6811	φ And	B6.5 IIIe	80	10387	9070	1980 Oct 17
23630	η Tau	B7 IIIe	140	10378	9060	1980 Oct 16
23862	28 Tau	B8 (V:)e-shell	320	10377	9059	1980 Oct 16
183914[b]	β Cyg B	B8 Ve	250	5369	4609	1979 May 26
195325	1 Del	B8–9:e-shell	280:	15514	11991	1981 Nov 17
88195	17 Sex	A1 V-shell	150	18819	14847	1982 Dec 19
42111[c]	HR 2174 (br)	A1 IV-shell	180	18821	14849	1982 Dec 19
112028[d]	HR 4893	A1 IV-shell	200	18822	14850	1982 Dec 19
38090	12 Lep	A3 III-shell	180	18820	14848	1982 Dec 19
39060	β Pic	A5 IV-shell	120	11216	9835	1981 Feb 1
192518	21 Vul	A5 IV-shell	200	15511	11988	1981 Nov 17
98058	φ Leo	A7 IV-shell	230	15504	11980	1981 Nov 15
108283	14 Com	F0 III-shell	200	15506	11982	1981 Nov 15

[a] HD 183656: *IUE* spectra obtained by Y. Kondo and J. Sahade.

[b] HD 183914: *IUE* spectra obtained by R. F. Wing.

[c] HD 42111: This is the brighter component of a visual binary system (ADS 4749; 5.8: 6.9; 29″ separation). The fainter component is a normal A0 V star with $v \sin i = 100$ km s^{-1} (Slettebak 1963).

[d] HD 112028: This is the brighter component of a visual binary system (ADS 8682; 5.3: 5.8; 22″ separation). The fainter component is itself a binary, showing a composite spectrum corresponding to two sharp-lined spectra of types A0 V and A2 V (Slettebak 1963).

TABLE 1B

THE STANDARD STARS

HD	Name/HR	Spectral Type	$v \sin i$ (km s^{-1})	SWP	LWR	Date of Observation
120315[a]	η UMa	B3 V	150	3170	2518	1978 Oct 1 and 27
34759	ρ Aur	B5 V	90	10389	9072	1980 Oct 17
219688	ψ^2 Aqr	B5 V	280	10385	9068	1980 Oct 17
87901	α Leo	B7 V	280	10379	9061	1980 Oct 16
222173	ι And	B8 V	90	10376	9057	1980 Oct 16
172167	α Lyr	A0 V	< 10	11213	9830	1981 Feb 1
97633	θ Leo	A2 V	15	15505	11981	1981 Nov 15
8538	δ Cas	A5 III–IV	100	11214	9831	1981 Feb 1

[a] HD 120315: *IUE* spectra obtained by A. Holm and C.-C. Wu.

stable levels and presumably are formed in an equatorial shell)—examples at various spectral types were selected and are listed in Table 1A. It should also be noted that while the Be stars all show emission, at least at Hα, the A–F type shell stars in Table 1A do not—they are recognized as shell stars because of the simultaneous presence of rotationally broadened lines, on the one hand, and sharp lines arising from ground states and metastable levels, on the other.

The presence of the B3e shell star 48 Lib in a list of Be stars of later types also deserves some explanation. This star has been classified as B5 in the literature, but this classification was probably based on the attribution of its rather strong Si II lines at 4128 and 4131 Å to the underlying star rather than to the shell. The correct spectral type appears to be close to B3, but we have retained 48 Lib as a program star because of its very interesting spectrum.

Table 1B lists the eight standard stars. These were selected to match, approximately, the spectral types of the program stars, and also to show a range of line broadening from sharp to broad. The spectral types are from Morgan, Abt, and Tapscott (1978), Morgan and Keenan (1973), or Johnson and Morgan (1953). The rotational velocities ($v \sin i$) are from Slettebak *et al.* (1975) or are recent determinations by us based on that system.

III. SUPERIONIZATION

An important result of ultraviolet observations of the hotter stars was the discovery in the spectra of some O- and B-type stars of atoms in higher stages of ionization than would be expected if normal photospheric radiation were the only ionizing source. Thus, lines of O VI, N V, C IV, and Si IV were observed in the spectra of stars with photospheres too cool to form these ions, calling for an additional source of ionization. This effect, usually called "superionization," has been reviewed recently by Marlborough (1982) and Doazan (1982).

a) Be Stars and A-Type Shell Stars

Examples of superionization in the Be stars have been reported by a number of investigators, including Doazan (1982), Henize *et al.* (1976), Lamers and Snow (1978), Marlborough (1977), Marlborough and Snow (1976), Morton (1976, 1979), and Panek and Savage (1976). Recently, using *IUE* spectra of 23 Be Stars with spectral types O9–B8 and luminosity classes III–V, Marlborough and Peters (1982) reported the probable presence of N V lines in Be stars as late as B3, C IV lines as late as B8 (B7, according to Slettebak 1982), and Si IV lines also as late as B8 (also B7, according to Slettebak 1982). They compare these observational results with theoretical estimates of the coolest spectral types (based on the work of Lamers and Snow 1978) at which each ion would be expected under radiative equilibrium conditions, either in the photosphere or in a cool circumstellar envelope: O7 for N V, B1 for C IV, and B3 for Si IV.

One object of this investigation was to determine whether superionization extends to spectral types later than B7 in the Be stars and, indeed, whether the phenomenon reaches even into the A-type shell stars. Figures 1–6 illustrate the Si IV resonance doublet (multiplet UV 1) at 1393.755 and 1402.770 Å, and the C IV resonance doublet (multiplet UV 1) at 1548.188 and 1550.762 Å in the spectra of a number of the program stars.

In Figures 1 and 2, both the Si IV and C IV lines are strong in absorption in the spectrum of the B3 shell star 48 Lib. Note that there is some distortion of the profiles by lines of Fe II from the shell. Figure 3 shows that the Si IV doublet is clearly present in the three B5–B6e stars shown there and is stronger in the two shell stars (ψ Per and o And) than in β Psc. The behavior of the C IV lines is consistent with that of Si IV: Figure 4 shows the lines clearly in the spectra of the same two B5–B6e shell stars but very weak or absent in β Psc.

Figures 5 and 6 show the Si IV and C IV doublets in some B6.5–B8e stars. The Si IV λ1393.755 line seems to be present in the three Be stars shown there, although other lines distort the spectrum somewhat. The C IV doublet also appears to be present in the spectra of η Tau and β Cyg B, but is very weak or absent in the B6.5e star ϕ And. The presence of both Si IV and C IV in the B8e star β Cyg B confirms the findings of Marlborough and Peters (1982) and Doazan (1982) that these lines persist into the late Be subtypes.

It is interesting that neither Si IV nor C IV is visible in the spectrum of the B8 shell star 28 Tau (Pleione). This

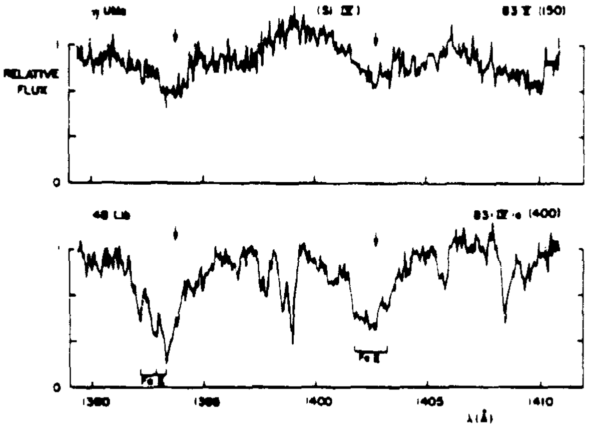

FIG. 1

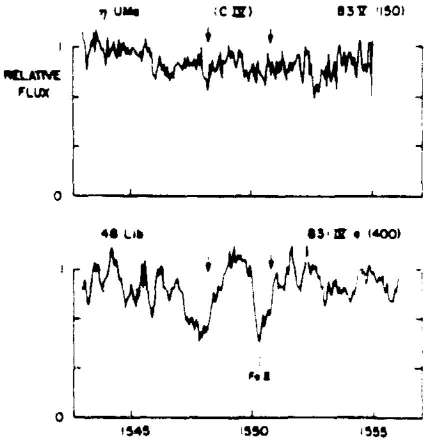

FIG. 2

FIG. 1.—Spectral region including the Si IV resonance doublet at 1393.755 and 1402.770 Å in the B3 shell star 48 Lib and the B3 V standard star η UMa. Rest wavelengths of the Si IV lines are indicated by arrows. The Fe II lines which distort the Si IV profiles in the spectrum of 48 Lib are at 1392.15, 1392.82, 1393.21, 1401.77, and 1403.25 Å.

FIG. 2.—Spectral region including the C IV resonance doublet at 1548.188 and 1550.762 Å in the B3 shell star 48 Lib and the B3 V standard star η UMa. Rest wavelengths of the C IV lines are indicated by arrows. An Fe II line at 1550.27 Å distorts one of the C IV profiles.

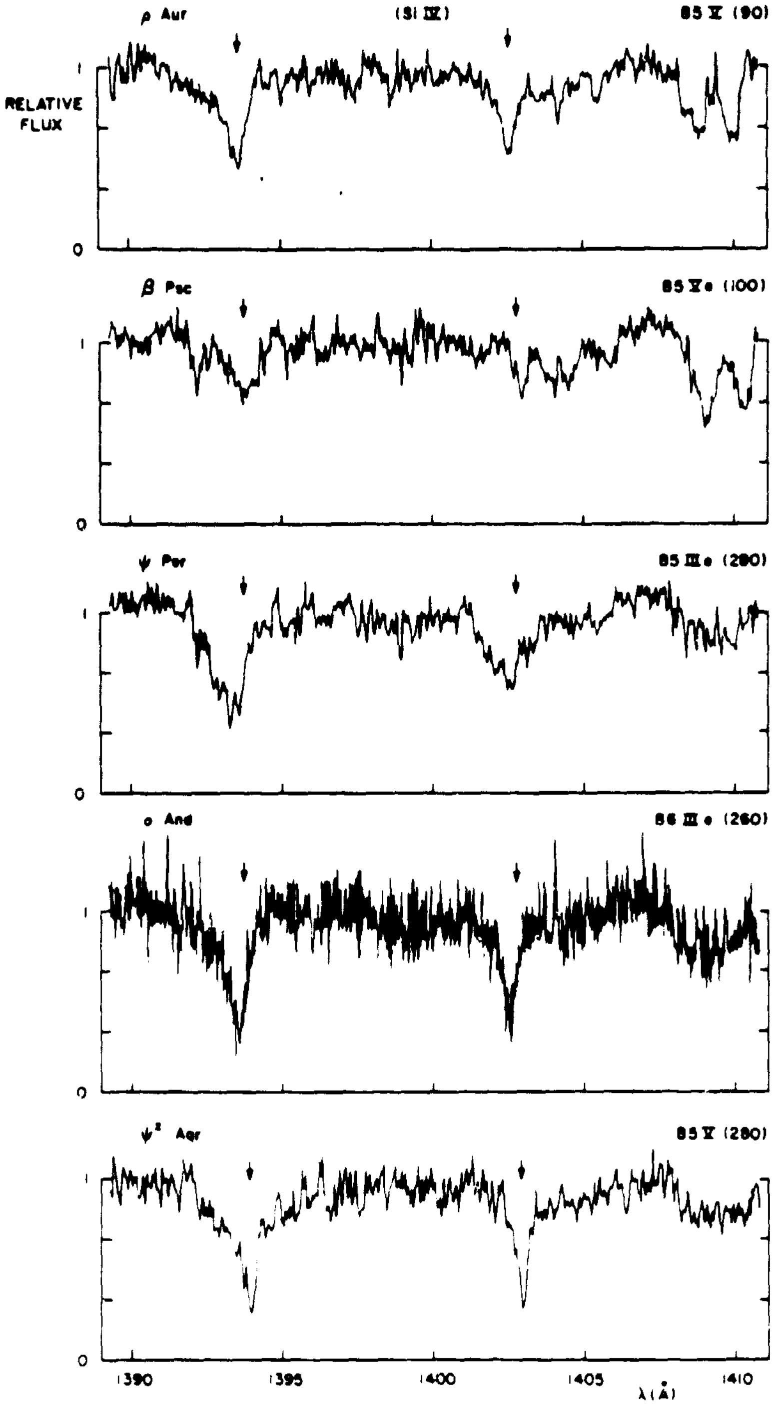

FIG. 3.—Spectral region including the Si IV resonance doublet at 1393.755 and 1402.770 Å (rest wavelengths indicated by arrows) in three B5–B6e stars and two B5 V standard stars.

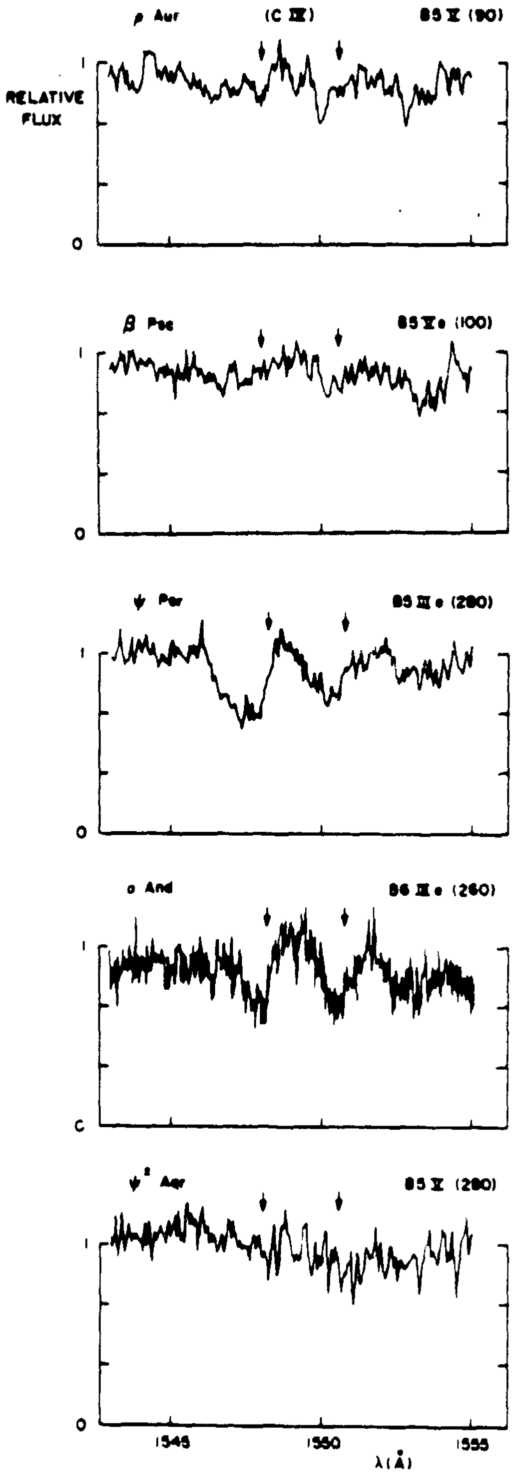

FIG. 4.—Spectral region including the C IV resonance doublet at 1548.188 and 1550.762 A (rest wavelengths indicated by arrows) in three B5–B6e stars and two B5 V standard stars.

object still has an extremely opaque shell, both in the visible and in the ultraviolet regions of the spectrum (see Figs. 16 and 21), which may affect the visibility of the Si IV and C IV lines. On the other hand, the late-type shell star 1 Del clearly shows Si IV λ1393.755 absorption. Its shell is not as dense as that of 28 Tau, but is sufficiently opaque in the visible portion of the spectrum to make the underlying star very difficult to classify. The estimated spectral type of 1 Del is B8–B9, making it one of the latest Be stars to show superionization effects.

Marlborough and Peters (1982) have noted some dependence of the strength of superionized lines in their B2e–B3e stars on *v* sin *i*. Visual inspection of their spectra revealed that the range of strengths of C IV lines depends on *v* sin *i*: while stars of large *v* sin *i* may have a range of C IV line strengths, stars of low *v* sin *i* seem to have only weak C IV lines. Although our sample is small, our spectra would seem to confirm this result. Inspection of Figures 3–6 shows that Be stars with small *v* sin *i* generally have weaker lines of Si IV and C IV than Be stars of the same spectral type with large *v* sin *i*. A possible interpretation, as Marlborough and Peters (1982) point out, is that the hot component of the circumstellar envelope in which the superionized lines arise is not spherically symmetrical.

Is superionization also a feature of A-type shell spectra? The answer seems to be no. There is no trace of the C IV resonance doublet in the spectra of our A–F type shell stars. The Si IV lines are also probably not present, although there is a dip in the A-type shell star spectra near the Si IV λ1393.755 line. Figure 7 shows this wavelength region for the A1 IV shell star HR 4893, with the A0 V standard α Lyr also shown for comparison. The Fe II lines at 1392.15, 1392.83, and 1393.21 Å are prominent, and probably account for the observed depression. This conclusion is strengthened by the fact that the other component of the Si IV doublet at 1402.769 Å appears not to be present. Unfortunately, our spectra are weak at these wavelengths, which makes the reality of broad features difficult to ascertain.

Of course, superionization may occur in the form of lines other than those of Si IV and C IV, and, indeed, in the A-type shell stars we might expect less ionized species than in the Be stars. In particular, if the Auger ionization mechanism proposed by Cassinelli and Olson (1979) is correct, the dominant stage of ionization in the stellar wind expected under radiative equilibrium conditions is doubly ionized by X-rays. Unfortunately, the resonance lines of C III, Si III, N II–IV, and O II–V are either outside the spectral range of *IUE* or too weak to detect in our *IUE* spectra, but we were able to examine the resonance lines of a number of other species:

1. Si II multiplet UV 2 (1533.445 and 1526.719 Å). These lines increase in strength through the B-type stars

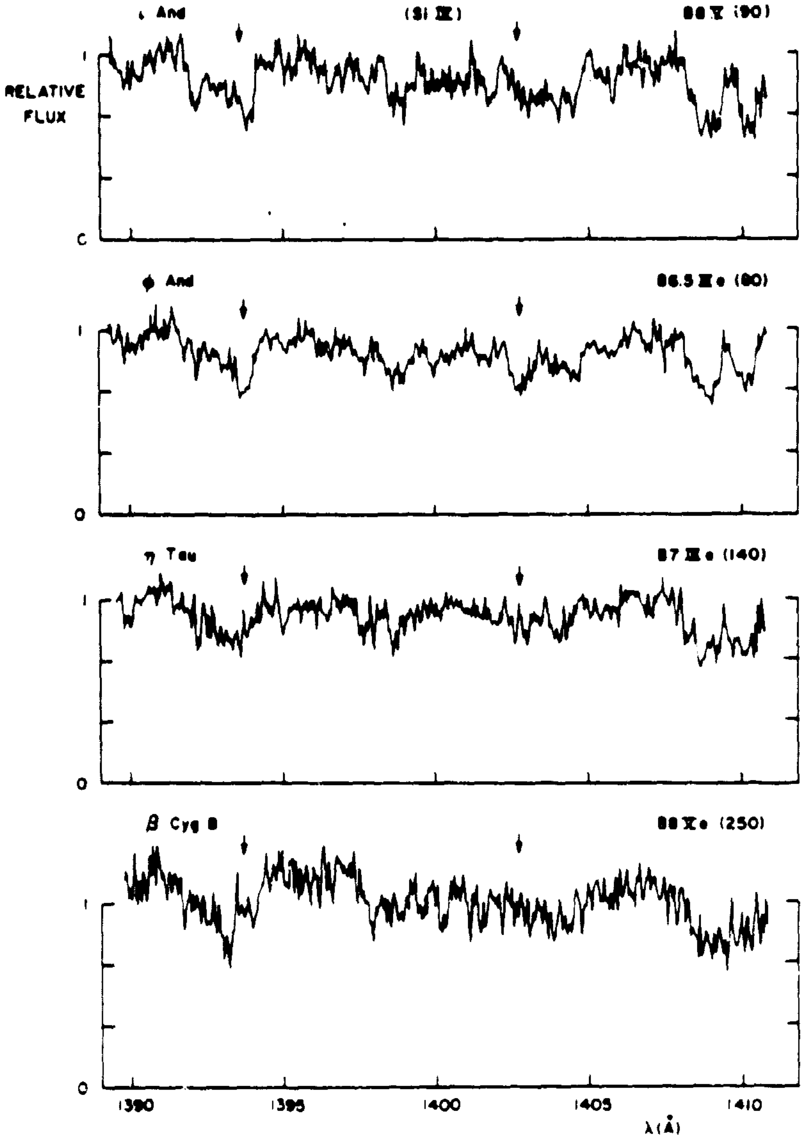

FIG. 5.—Spectral region including the Si IV resonance doublet at 1393.755 and 1402.770 Å (rest wavelengths indicated by arrows) in three B6.5–B8e stars and a B8 V standard star.

and are very strong at A2, with no significant difference between A-type shell stars and A-type standard stars. Other *IUE* spectra are too weak to show the lines for types later than A2.

2. C II multiplet UV 1 (1335.708 and 1334.532 Å). Our *IUE* spectra show these lines to strengthen from the middle to the late B types and remain strong through A2 (again, our sensitivity is too low for detection later than A2), with no appreciable difference between shell-type and normal stars.

3. Al II multiplet UV 2 (1670.787 Å). This line is weak at B3, strengthens through the B types, and becomes very strong in the range B8–A5 (A5 being the limit of our detection). The line is shown in A0–A5 stars in Figures 8 and 9, where again no significant difference between shell-stars and standard stars is apparent.

4. Al III multiplet UV 1 (1854.716 and 1862.790 Å). These lines are strong in the middle B-type stars, as shown in Figure 10, which also shows them to be somewhat stronger and asymmetrical in the two shell stars (ψ Per and o And), indicating contributions from a stellar wind. This will be discussed in the next section. The Al III lines exhibit a general weakening from B8 into the late A types, with no striking differences be-

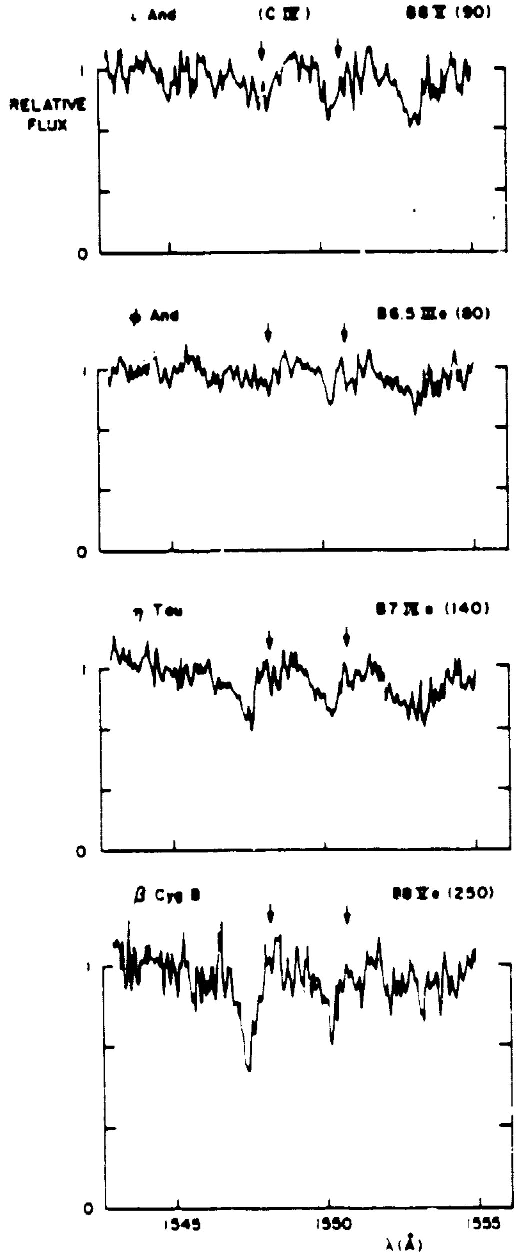

FIG. 6.—Spectral region including the C IV resonance doublet at 1548.188 and 1550.762 Å (rest wavelengths indicated by arrows) in three B6.5–B8e stars and a B8 V standard star.

tween the A shell and the A-type standard stars. The lone exception is β Pic (A5 IV shell), which has rather strong Al III lines relative to standard stars.

In summary, the various resonance lines we examined in our *IUE* spectra show no general tendency toward superionization in the A-type shell stars.

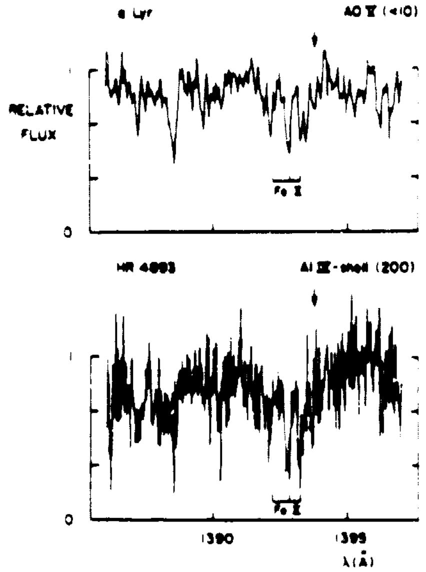

FIG. 7.—Spectral region near the Si IV resonance line at 1393.755 Å in an A1 shell star and the A0 V standard α Lyr (rest wavelength indicated by an arrow). The dip near the Si IV line appears to be due to the Fe II lines at 1392.15, 1392.82, and 1393.21 Å.

b) *The Standard Stars*

An interesting result of this investigation is that some, but not all, of the standard, non-emission-line stars also show superionization effects in their ultraviolet spectra. Figures 1 and 2 show that the B3 V standard η UMa has Si IV but not C IV absorption, as might be expected at this spectral type. However, both the sharp-lined (ρ Aur) and broad-lined (ψ^2 Aqr) B5 V standards show strong Si IV absorption (Fig. 3), while the C IV doublet is weak or absent in both (Fig. 4).

For later spectral types it is very difficult to be certain whether superionized species exist. The broad-lined B7 V standard α Leo appears to show a broad and shallow Si IV λ1393.755 line, but the other component of the doublet is difficult to see. Figure 5 shows the depression at the position of the Si IV λ1393.755 line in the spectrum of ι And, but it is distorted by the presence of a number of Fe II lines. Si IV λ1402.769 seems not to be present. The C IV resonance doublet is not detectable in the spectrum of α Leo but may be present in ι And (see Fig. 6). Moving to still later types, Figure 7 shows a depression near Si IV λ1393.755 in the spectrum of the A0 V standard α Lyr, but this is probably due entirely to the confluence of the several Fe II lines in that wavelength region. There is no trace of the C IV resonance doublet, nor does the A2 V standard θ Leo show either Si IV of C IV absorption.

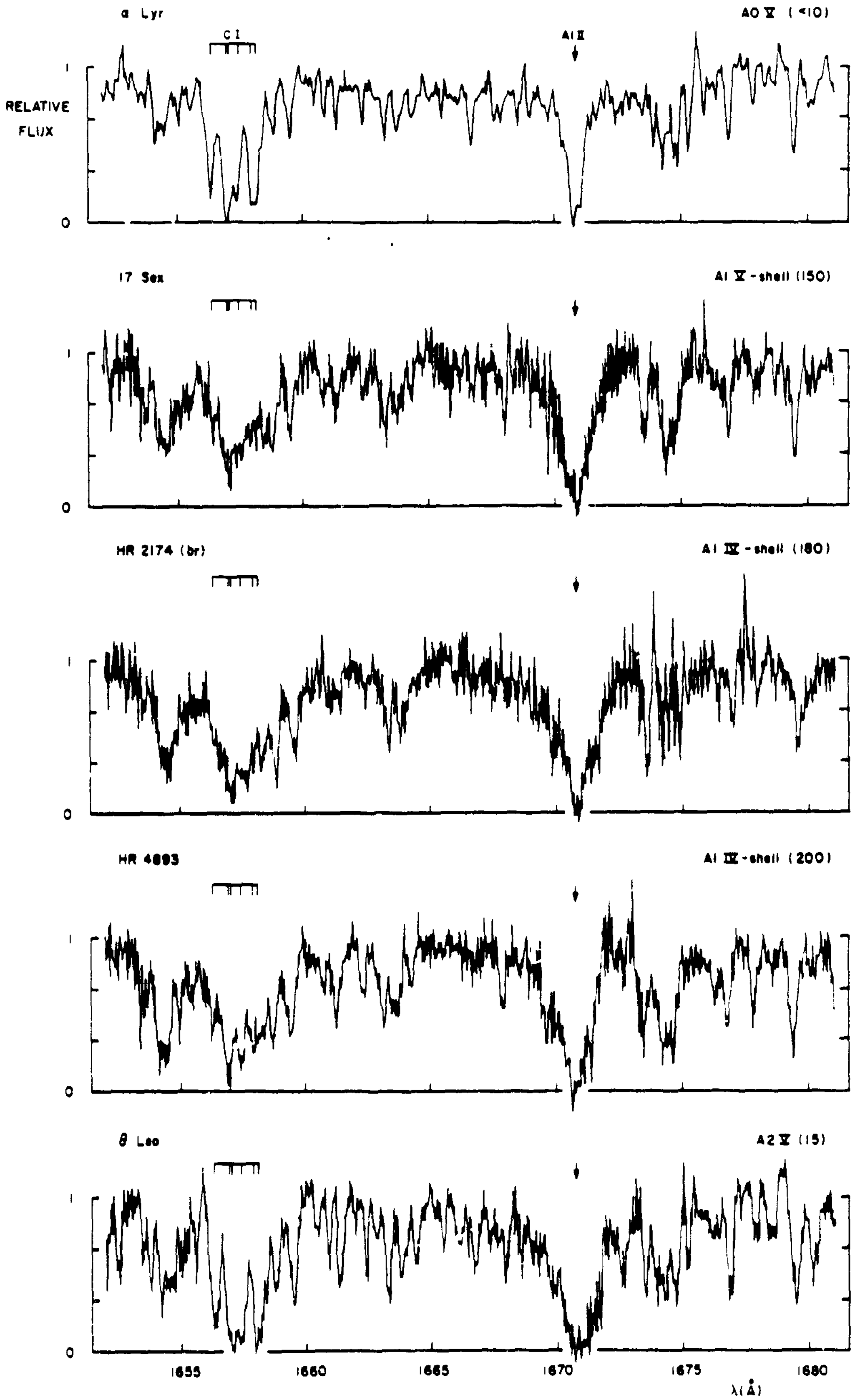

FIG. 8.—Wavelength region including the C I resonance multiplet (UV 2) centered at 1657 Å and the Al II resonance line (UV 2) at 1670.787 Å in three A1 shell stars and an A0 V and A2 V standard star.

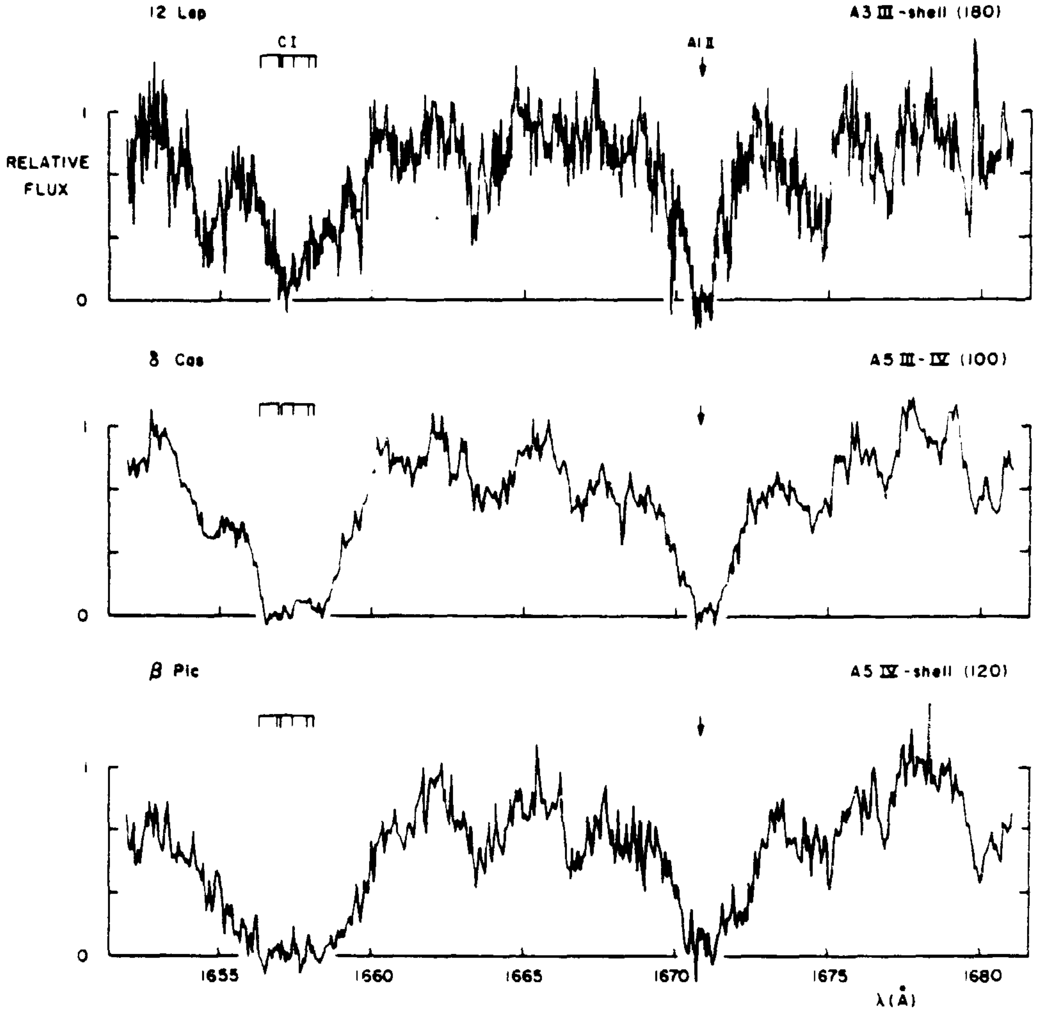

FIG. 9.—Wavelength region including the C I resonance multiplet (UV 2) centered at 1657 Å and the Al II resonance line (UV 2) at 1670.787 Å in two A3–A5 shell stars and an A5 III–IV standard star.

We have already discussed resonance lines of other species in the last section and found no strong evidence for unusual line strengths in either the program or standard stars of later types. Our observations thus show that superionization exists in standard, non-emission-line stars as late as B5 and, possibly, as late as B7–B8, but probably not later.

IV. STELLAR WINDS AND MASS LOSS

There is now a considerable literature on stellar winds and mass loss from Be stars. Marlborough (1982) and Doazan (1982) have summarized the ultraviolet data which have been used to infer or demonstrate mass loss in Be stars. Much of the work has centered on Be stars earlier than B2 (e.g., γ Cas, 59 Cyg, ζ Tau, ϕ Per). We note here two papers which include also Be stars of somewhat later type: Dachs's (1980) *IUE* observation of eight bright Be stars, and the work of Snow (1981) based on *Copernicus* scans. The latter determined mass-loss rates for 19 Be stars from analysis of Si III and Si IV resonance lines, including objects as late as spectral type as B6.

Asymmetrical or violet-displaced resonance absorption lines, suggesting mass loss, were observed in all (except 28 Tau) of our Be program stars and also in a number of our B-type standard stars, but not in the A–F type shell stars. This is shown for the Si IV and C IV doublets in Figures 1–6. In addition to these lines, stellar winds are also suggested in the profiles of resonance lines of Si II (multiplet UV 5), Al III (multiplet UV 1), and Mg II (multiplet UV 1) in some of our stars. Figure 10 shows the Al III resonance doublet at 1854.716 and 1862.790 Å in the spectra of three B5–B6e stars and two B5 standard stars. Note the asymmetrical structure in the shell stars ψ Per and o And, as well as in the broad-lined standard ψ^2 Aqr. On the other hand, the relatively sharp-lined (pole-on) Be star η Tau also shows a slight asymmetry in the same sense.

The standard B5 V star ρ Aur also shows asymmetrical resonance lines of other ions, including Si II and Mg II. The Mg II resonance doublet at 2795.528 and 2802.704 Å in the spectra of several Be and standard stars is shown in Figure 11, where the lines are clearly seen to be asymmetrical in the spectrum of ρ Aur.

Attempts to determine quantitatively the rates of mass loss associated with the winds from Be stars from ultraviolet line profiles have been summarized by Snow (1982*a*). A number of investigators, including Snow (1981), have utilized the atlas of theoretical P Cygni and stellar wind profiles by Castor and Lamers (1979). The latter assume resonance scattering and use the Sobolev approximation. They characterize the expanding envelope by two functions: the velocity law $v(r)$ and the optical depth $\tau(v)$. Thus, if v_∞ is the terminal velocity and R_* the photospheric radius, they define

$$w = v/v_\infty, \qquad x = r/R_*,$$

and choose a velocity law of the type

$$w = 0.01 + 0.99(1 - 1/x)^\beta \qquad (\beta > 0).$$

For an optical depth law Castor and Lamers adopt, for the radial optical depth $\tau_{\rm rad}(w)$, a law giving zero residual intensity at the violet edge of the line profile (i.e., at $w = -1$):

$$\tau_{\rm rad}(w) = T(\gamma + 1)(1 - w_0)^{-1-\gamma}(1 - w)^\gamma \qquad (\gamma \geq 0),$$

where

$$T = \frac{\pi e^2}{mc}\frac{f\lambda_0}{v_\infty}N_i,$$

f is the absorption oscillator strength, λ_0 is the rest wavelength of the transition, and N_i the column density of the absorbing ion in the envelope. Castor and Lamers take $w_0 = 0.01$ for the flow speed at the base of the envelope, choose $\beta = 0.5$, 1, 2, and 4, and point out that the absorption part of their P Cygni profiles is very insensitive to the velocity law. Values of γ range between 0.5 and 4. They then give theoretical stellar wind profiles for a variety of combinations of β, γ, and T.

Since the lines being studied may be formed in the photosphere as well as in the winds of the stars under consideration, it is necessary to correct for the presence of underlying photospheric profiles. Castor and Lamers suggest a procedure for this correction, which requires that the photospheric profile be known. We follow their

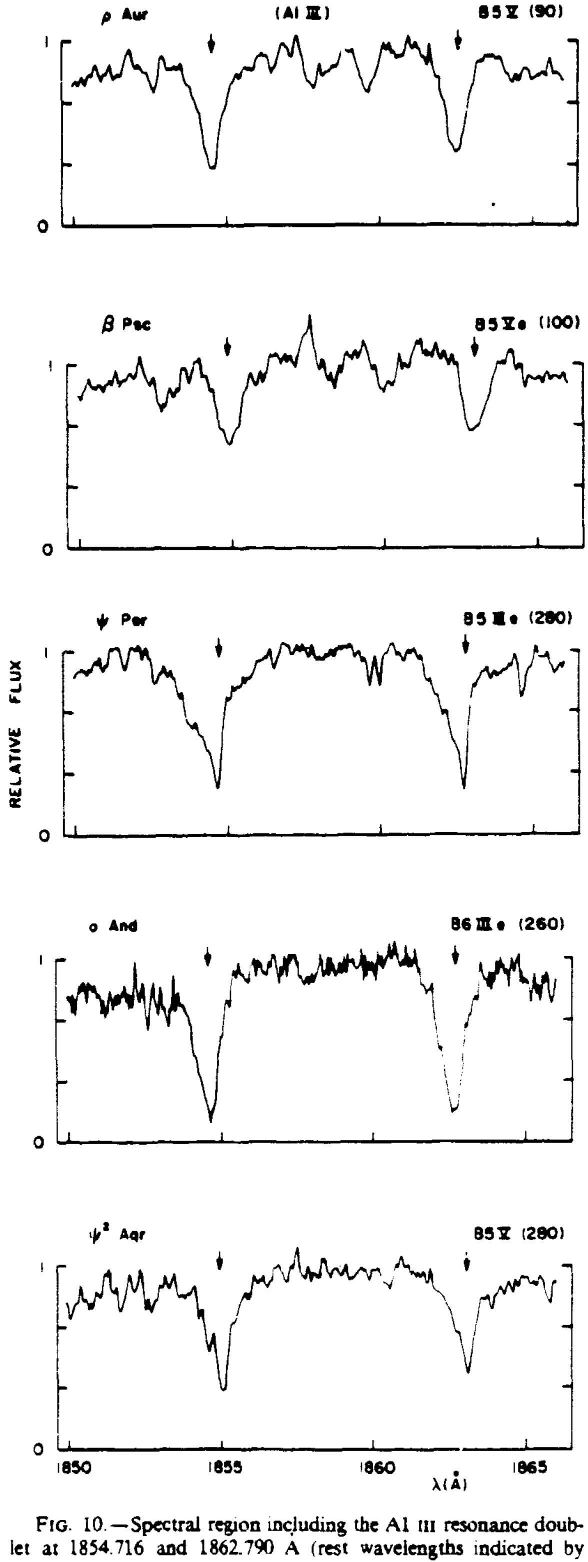

FIG. 10.—Spectral region including the Al III resonance doublet at 1854.716 and 1862.790 Å (rest wavelengths indicated by arrows) in three B5–B6e stars and two B5 V standard stars.

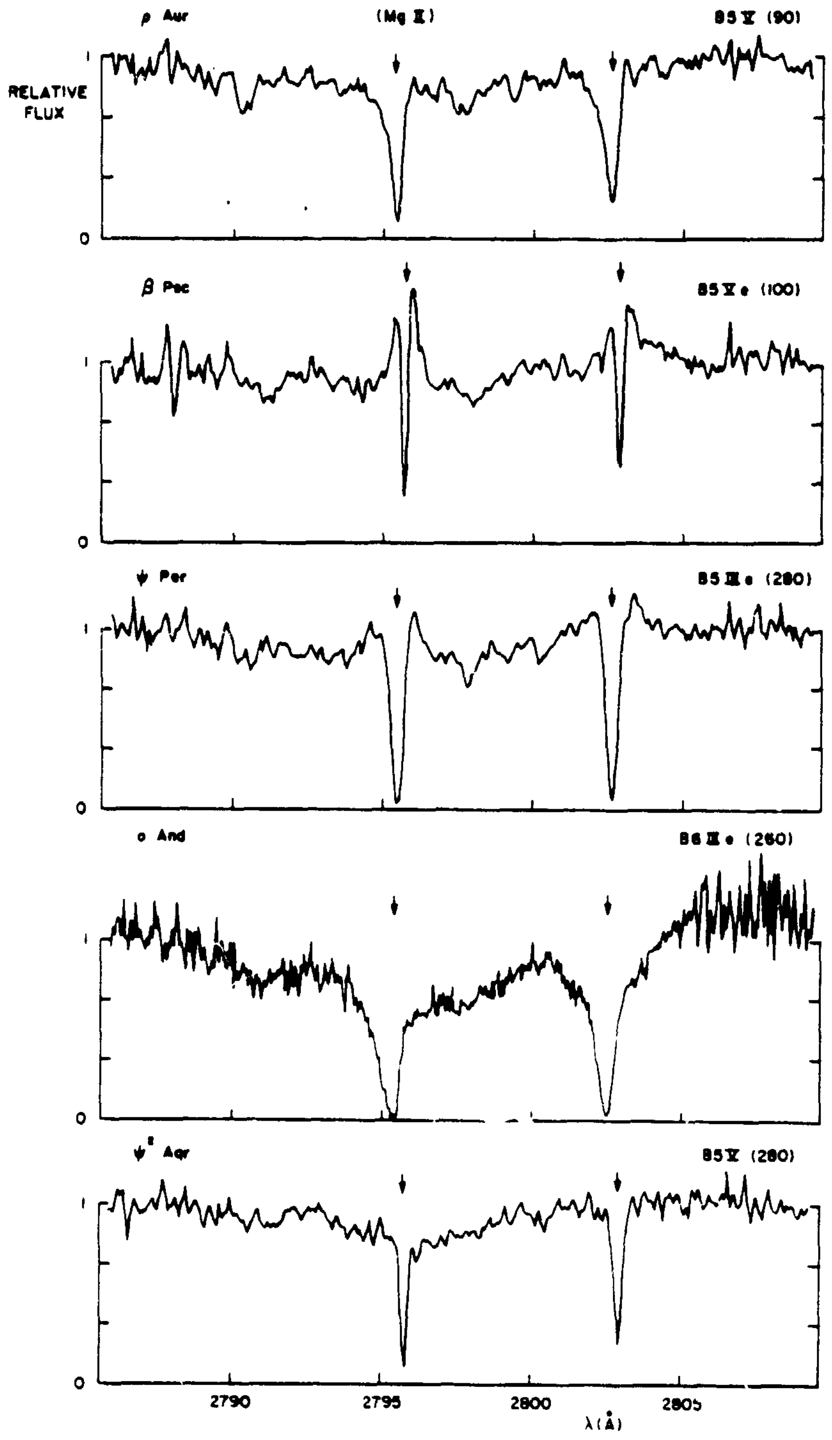

FIG. 11.—Spectral region including the Mg II resonance doublet at 2795.528 and 2802.704 Å (rest wavelengths indicated by arrows) in three B5–B6e stars and two B5 V standard stars.

TABLE 2

THE ATOMIC DATA

Ion	A_E ($H = 1.0$)	λ_0 (A)	f	IP (eV)
Si IV	4.5×10^{-5}	1393.755 1402.770	0.528 0.262	45.14
Si II......	4.5×10^{-5}	1190.416	0.251	16.35
C IV	4.2×10^{-4}	1548.188 1550.762	0.194 0.097	64.49
Al III ...	3.3×10^{-6}	1854.716 1862.790	0.539 0.268	28.33
Mg II ...	4.0×10^{-5}	2795.528 2802.704	0.592 0.295	14.97

method here and adopt also the procedure by Snow (1981), who derived the photospheric profiles directly from the observational data. He assumed that the long-wavelength side of a photospheric line profile will be virtually unaffected by the wind if it is optically thin, and folded the long-wavelength side of his observed profiles over to obtain an approximation to the underlying photospheric profile.

Still following Snow's (1981) application of the Castor and Lamers (1979) theoretical profiles, we calculated our mass-loss rates from

$$\dot{M} = 1.74\times10^{-18}\tau(w)\frac{v_\infty^2 R_*}{f\lambda_0 g A_E},$$

where $\tau(w)$ is calculated for $w = 0.5$, v_∞ is in units of km s^{-1}, R_* is in solar radii, λ_0 is in angstroms, g is the ionization fraction for the observed species, A_E is the assumed abundance of the observed element with respect to hydrogen, and the mass-loss rate is computed in solar masses per year.

Table 2 lists the atomic data for the lines analyzed in this investigation. The element abundances A_E are solar values from the compilation by Ross and Aller (1976), while the wavelengths and f-values are from Morton (1978).

There are many uncertainties in the computation of mass-loss rates from line profiles, some of which we will discuss later in this section. The greatest uncertainty is probably the estimation of the ionization fraction g. This is because we have no complete physical picture of how superionization arises; collision and radiative processes may both play a role (see the discussion and references in Marlborough 1982 and Snow 1981). In the absence of definitive ionization equilibrium models, we will assume that *all* of a particular atomic species exists in the form of the ion being observed; i.e., we will take $g = 1$ in all cases. This means that our mass-loss rates will be *lower limits* in every case, since the atom may exist in other stages of ionization which could also contribute to the mass flow.

Table 3 lists the computed lower limits to the mass-loss rates for both the program stars and those standard stars which show asymmetrical or violet-shifted resonance line profiles. The stars are listed approximately in order of spectral type, from the hottest to the coolest. The Castor and Lamers (1979) parameter β has the value of 0.5 for all cases and is not listed in Table 3 with their other parameters γ and T. The stellar radii were estimated from the papers by Code *et al.* (1976) and Harris, Strand, and Worley (1963).

In addition to the unknown ionization balance in the stellar winds, other uncertainties affect the computation of mass-loss rates. These have been pointed out by Snow (1981) and, in a different context, by Kurucz (1974). One difficulty lies in the determination of the terminal velocity v_∞, which has considerable uncertainty because the short-wavelength wing of the line profile merges very gradually into the continuum. Snow (1981) points out, however, that the error in estimating v_∞ is partially cancelled out in the profile fitting process, with the mass-loss rate not affected radically. Another possible source of error is the blending of the short-wavelength wing of Si IV λ1393.755 with lines of Fe II, as can be seen in Figures 1 and 7. This effect would be especially troublesome for the shell stars and may lead to values of the terminal velocity which are too large and therefore also lead to excessive mass-loss rates. The assumption of spherical symmetry in the Castor and Lamers (1979) analysis may also not be valid for winds from Be stars, although the evidence is somewhat mixed at this time (see references in Snow 1981). Kurucz (1974), in a sample spectral synthesis near the C IV resonance doublet for a B2 star, has pointed out that the continuum level and equivalent widths will be affected by rotational line broadening. Furthermore, he shows that the central wavelengths of the rotated features are a function both of the blending with many weaker lines and the rotational velocity. Wavelength shifts produced in this way might then be erroneously interpreted as Doppler shifts due to mass motion in the stellar atmosphere.

Recognizing that our assumptions are the same as his, and that there is evidence (see Marlborough 1982; Doazan 1982) that Be winds are variable with time, it is nevertheless interesting to compare our results with those of Snow (1981). For three stars in common (48 Lib, ψ Per, and o And), our terminal velocities are within 25% of his, and our mass-loss rates derived from the Si IV λ1393.755 line (adopting his ionization fractions) agree with his to within a factor of 2. On the other hand, Table 3 shows that our mass-loss rates as derived from the C IV lines are systematically lower and those from the Al III lines systematically higher than the Si IV mass-loss rates for a given star. A partial explanation may lie in the unknown ionization equilibria for all three elements. Also, we have already mentioned that possible blending of Fe II lines with the short-wavelength wing of

TABLE 3

TERMINAL VELOCITIES AND LOWER LIMITS TO THE MASS-LOSS RATES

Star	$R_*/R_\odot$	M_{bol}	T_e (K)	Line	v_∞ (km s^{-1})	γ	T	Lower Limit to $\dot{M}$ ($M_\odot$ yr^{-1})
48 Lib	6	−4.1	18,000	Si IV λ1393.755	610	2	0.25	2.2×10^{-11}
				C IV λ1548.188	490	1	0.25	5.2×10^{-12}
ψ Per	6	−3.5	15,000	Si IV λ1393.755	530	2	0.5	3.5×10^{-11}
				C IV λ1548.188	450	1	0.5	8.5×10^{-12}
				Al III λ1854.716	420	2	0.5	2.2×10^{-10}
β Psc	4	−2.4	15,300	Si IV λ1393.755	480	0.5	0.1	5.3×10^{-12}
ρ Aur	4	−2.4	15,300	Si IV λ1393.755	700	1	0.25	2.7×10^{-11}
				Si II λ1190.416	200	2	1	1.6×10^{-11}
				Mg II λ2802.704	230	2	0.25	2.1×10^{-12}
ψ^2 Aqr......	4	−2.4	15,300	Si IV λ1393.755	480	1	0.1	4.8×10^{-12}
				Al III λ1854.716	320	1	0.1	2.2×10^{-11}
o And.......	5	−3.1	14,000	Si IV λ1393.755	740	2	0.25	2.7×10^{-11}
				C IV λ1548.188	410	0.5	0.25	3.1×10^{-12}
				Al III λ1854.716	380	2	0.25	7.2×10^{-11}
HR 7415 ...	5	−2.5	14,000	Si IV λ1393.755	600	0.5	0.25	2.6×10^{-11}
				C IV λ1548.188	300	0.5	0.1	6.8×10^{-13}
φ And	5	−2.9	13,700	Si IV λ1393.755	570	0.5	0.1	9.4×10^{-12}
η Tau	5	−2.7	13,200	Si IV λ1393.755	550	0.5	0.1	8.7×10^{-12}
				C IV λ1548.188	450	0.5	0.25	3.8×10^{-12}
				Al III λ1862.790	230	1	0.1	2.8×10^{-11}
α Leo	3.6	−1.6	13,400	Si IV λ1393.755	530	0.5	0.1	5.8×10^{-12}
β Cyg B	3	−0.7	12,000	Si IV λ1393.755	590	0.5	0.1	6.0×10^{-12}
				C IV λ1548.188	260	1	0.5	1.4×10^{-12}
1 Del	3	−1.0	11,500	Si IV λ1393.755	600	0.5	0.25	1.5×10^{-11}

Si IV λ1393.755 may have resulted in systematically large Si IV mass-loss rates.

Based on computed mass-loss rates, Snow (1982*b*) found the three non-Be stars in his sample to be indistinguishable from the Be stars and suggested that the Be phenomenon itself is not uniquely linked to the presence of stellar winds. Although our mass-loss rates in Table 3 may be quite uncertain, there is some validity in comparing the stars listed among themselves, as Snow has done. We then also find no striking differences between the Be stars and the standard stars, with one exception: the five Be shell stars as a group appear to have somewhat larger mass-loss rate lower limits than the other stars in the table. This is also shown in Figures 12, 13, and 14. Obviously, more observations of both normal stars and Be stars are necessary to see whether this correlation is valid.

The terminal velocities v_∞ listed in Table 3 range between 200 and 740 km s^{-1}, somewhat lower than either the values measured by Dachs (1980) or those listed by Marlborough (1982) for hotter Be stars. Our

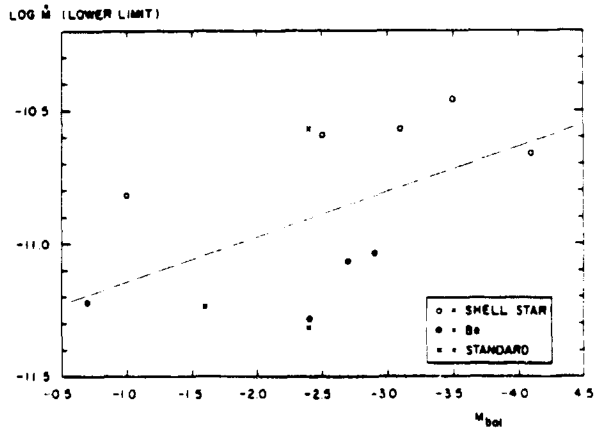

FIG. 12.—Lower limits to the mass-loss rates derived from the Si IV λ1393.755 line for 12 Be and standard stars vs. their bolometric absolute magnitudes. Dashed line represents a least-squares fit to the data.

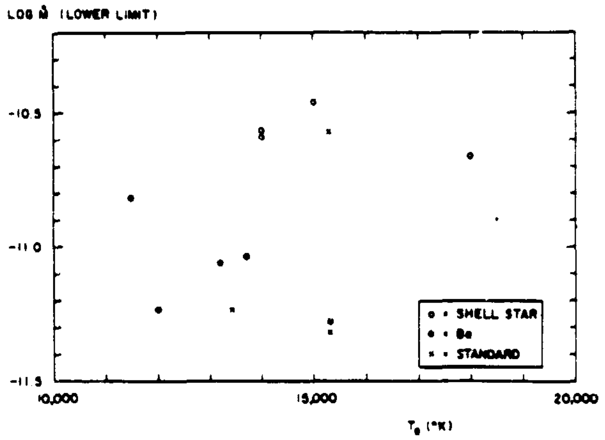

FIG. 13.—Lower limits to the mass-loss rates derived from the Si IV $\lambda 1393.755$ line for 12 Be and standard stars vs. their effective temperatures.

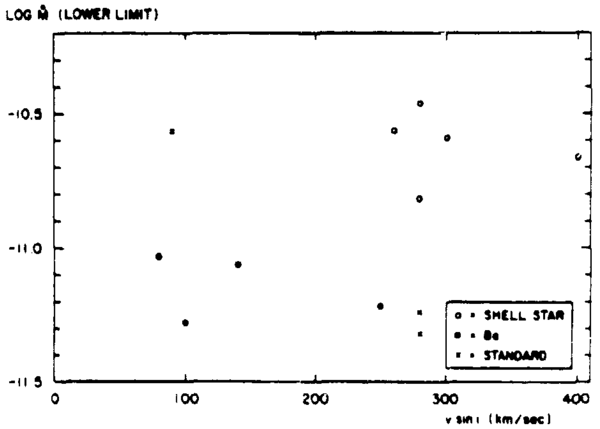

FIG. 14.—Lower limits to the mass-loss rates derived from the Si IV $\lambda 1393.755$ line for 12 Be and standard stars vs. their rotational velocities ($v \sin i$).

measured terminal velocities show no clear correlation with ionization potential. In general, v_∞ is largest for Si IV, followed by C IV, Al III, Mg II, and Si II. Three Be stars in particular (ψ Per, o And, and η Tau) show simultaneously asymmetrical resonance lines of Si IV, C IV, and Al III, with terminal velocities decreasing in that order. This result is not consistent with the usual assumption that ionization increases outward, in which case the C IV lines should show the largest terminal velocities. Again, however, the Si IV terminal velocities could be systematically high due to blending of Fe II lines with the short-wavelength wing of Si IV $\lambda 1393.755$.

We may attempt to correlate our mass-loss rates with various stellar parameters. Snow (1982*b*) has discussed the extension of OB star winds to lower luminosities, finding that luminosity is the dominant factor in driving the mass loss from Of to mid-B main-sequence stars. Specifically, he believes that the winds in Be stars may represent a straightforward extension of the OB star wind phenomenon to lower luminosities. Insofar as we find stellar winds in the Be stars of later type but not in the A–F shell stars, which are of lower luminosity than the Be stars, our results are consistent with Snow's suggestion.

Is there a correlation between the mass-loss rates and the luminosities of our B-type stars? Figure 12 shows our lower limits to the mass-loss rates derived from the Si IV $\lambda 1393.755$ line for 12 stars versus their bolometric absolute magnitudes. The latter were estimated from the calibrations by Blaauw (1963) and Keenan (1963) and the bolometric corrections of Code *et al.* (1976) and Flower (1977), and are listed in Table 3. Again, a correlation seems to be present in the sense of Snow's (1982*b*) conclusion: the more luminous stars show somewhat larger mass-loss lower limits. The dashed line in Figure 12 represents a least-squares fit to the data and has a correlation coefficient $r = 0.50$.

A correlation between mass loss and effective temperatures also exists in the sense that mass-loss effects are found in the B-type stars but not in the A-type stars. Within the B-type stars which show mass-loss effects, however, no correlation with effective temperatures is evident. This is shown in Figure 13, in which our mass-loss lower limits derived from the Si IV $\lambda 1393.755$ line are plotted versus effective temperatures for 12 stars. The latter were estimated from the papers by Code *et al.* (1976) and Flower (1977), and are listed in Table 3. A straight-line least-squares fit to the data in Figure 13 has a correlation coefficient $r = 0.26$, which we consider not to be significant.

Figure 14 shows our lower limits to the mass-loss rates derived from the Si IV $\lambda 1393.755$ line for the same 12 stars plotted against their rotational velocities $v \sin i$. Here we might expect a correlation for the Be stars if stellar winds are related to the Be phenomenon, since the latter is clearly related to stellar rotation. No correlation is evident: a straight-line least-squares fit to the data in Figure 14 has a correlation coefficient r of only 0.26.

Plots of bolometric absolute magnitudes, effective temperatures, and rotational velocities for the B-type stars versus lower limits to their mass-loss rates derived from C IV and Al III lines were also made. Although fewer observations were involved, these generally showed the same trends as the Si IV plots: no correlations of mass loss with effective temperatures and rotational velocities, but a mild correlation with bolometric absolute magnitudes.

V. EMISSION LINES

Be stars are characterized by Balmer emission in the visible portion of the spectrum, but show little or no emission in the ultraviolet. In a study of eight bright Be stars, selected because their optical spectra show very strong Balmer line emission, Dachs (1980) found emission in *IUE* spectra only in the wings of the Mg II resonance doublet and (for two stars only) in the wings of the Fe II resonance multiplets UV 1 and UV 2. Bruhweiler, Morgan, and van der Hucht (1982) observed Mg II *h* and *k* emission in the spectra of the early-type Be stars ϕ Per, υ Cyg, and γ Cas using the balloon-borne ultraviolet stellar spectrometer (BUSS), and suggested that a Bowen mechanism driven by Ly β is responsible.

Only two of our program stars show definite emission in the wings of the Mg II resonance lines: β Psc and ψ Per. These lines are shown in Figure 11. It is interesting that both are B5 stars, but ψ Per is a shell star with a large $v \sin i$, whereas ψ Psc has a small $v \sin i$ and is presumably a pole-on star. The Mg II resonance-line region in the spectrum of the B3 shell star 48 Lib is shown in Figure 15, along with the corresponding wavelength region in the standard B3 V star η UMa. Notice that the Mg II lines in 48 Lib are deep and show velocity structure, suggesting motions in the shell but no obvious emission. Dachs (1980) also found Mg II emission in 48 Lib to be uncertain.

In order to check for possible variability of the Mg II emission features in the spectra of the aforementioned stars, additional *IUE* spectra of ψ Per (LWR 9832 on 1981 February 1 and LWR 11990 on 1981 November 17) and β Psc (LWR 11989 on 1981 November 17) were obtained. No significant differences are discernable in these spectra, indicating no major changes in the Mg II emission over time periods of the order of a few months to a year for these stars.

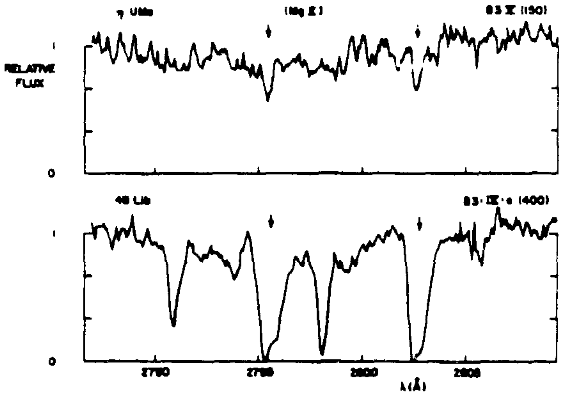

FIG. 15.—Spectral region including the Mg II resonance doublet at 2795.528 and 2802.704 A (rest wavelengths indicated by arrows) in the B3 shell star 48 Lib and the B3 V standard star η UMa.

The Mg II resonance doublet in the spectra of other program and standard stars is shown in Figures 16, 17, and 18. There is no obvious emission for any of the stars, and we conclude that Mg II emission does not occur in spectral types later than B5 in our sample.

It is interesting that Blanco *et al.* (1982) find faint chromospheric emission in the Mg II *k* line of α Aql, a rapidly rotating A7 which is on or near the main sequence. On the other hand, they find no Mg II emission in HD 192518 = 21 Vul, which is one of our A-type shell program stars. We also see no Mg II emission in the spectrum of 21 Vul (see Fig. 18), nor is emission present in any of our other A–F type shell stars.

Figures 19, 20, and 21 show the wavelength region which includes the strongest lines of the Fe II multiplet UV 1 in some of our Be stars and standard stars. Figures 22 and 23 show lines of the multiplets UV 62, UV 63, and UV 64 of Fe II in the spectra of our A–F type shell stars plus standard stars. The only object which may have faint Fe II emission wings is the B5 shell star ψ Per (see Fig. 20, in which the short-wavelength wings of the Fe II multiplet UV 1 lines appear to be somewhat higher than the long-wavelength wings), but this is uncertain.

VI. THE A–F TYPE SHELL STARS

The eight A–F type shell stars listed in Table 1A range in spectral types from A1 to F0. In the optical spectrum, all show rotationally broadened lines ($v \sin i$ ranging from 120 to 230 km s^{-1}) plus sharp lines which arise from ground states or metastable levels, the latter presumably formed in some kind of shell around the rapidly rotating underlying star. None of the stars show Balmer emission, although the hotter ones may have sharp Balmer absorption cores, which presumably also form in the shell. While the hotter A-type shell stars may also have sharp Fe II lines in the optical spectrum (arising from metastable levels 2.6 to 2.8 eV above the ground state), the cooler ones are more likely to show the Ti II lines at 3685, 3759, and 3761 Å (arising from metastable levels ~ 0.6 eV above the ground state) and the H and K resonance lines of Ca II.

We have already noted that the A–F type shell stars seem not to show superionized lines, nor do they have asymmetrical or violet-displaced resonance lines suggesting mass loss. Figures 22 and 23 show the strongest lines of Fe II multiplets UV 62, UV 63, and UV 64 (with lower metastable energy levels at ~ 1 eV) in the spectra of A–F type shell stars and standard stars. In Figures 8 and 9 we show the wavelength region around the Al II resonance line at 1670.787 Å, including also the resonance multiplet of C I centered at 1657 Å, for five A-type shell stars plus three standard stars. We have already noted that the A–F type shell stars seem not to

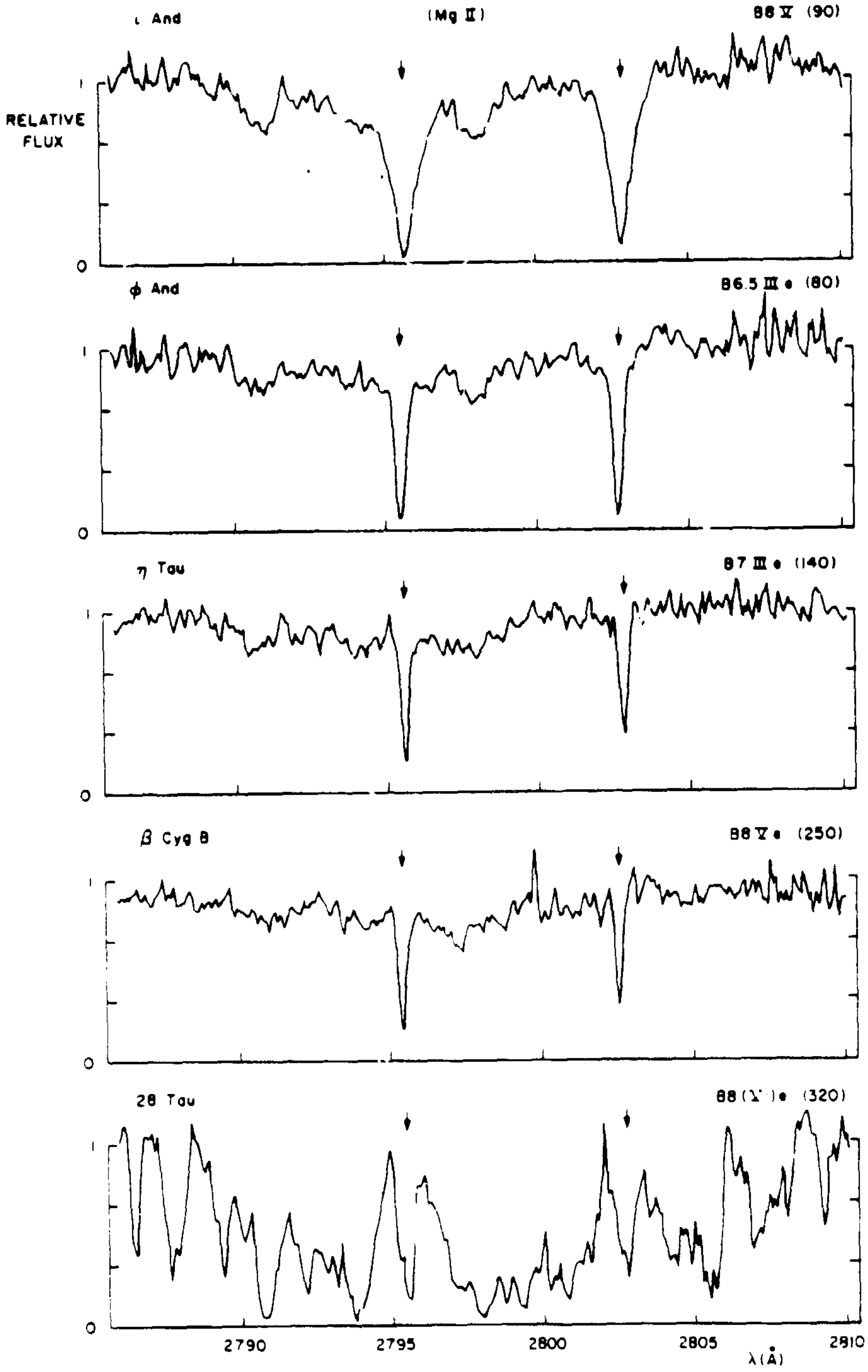

FIG. 16.—Spectral region including the Mg II resonance doublet at 2795.528 and 2802.704 Å (rest wavelengths indicated by arrows) in four B6.5–B8e stars and an B8 V standard star.

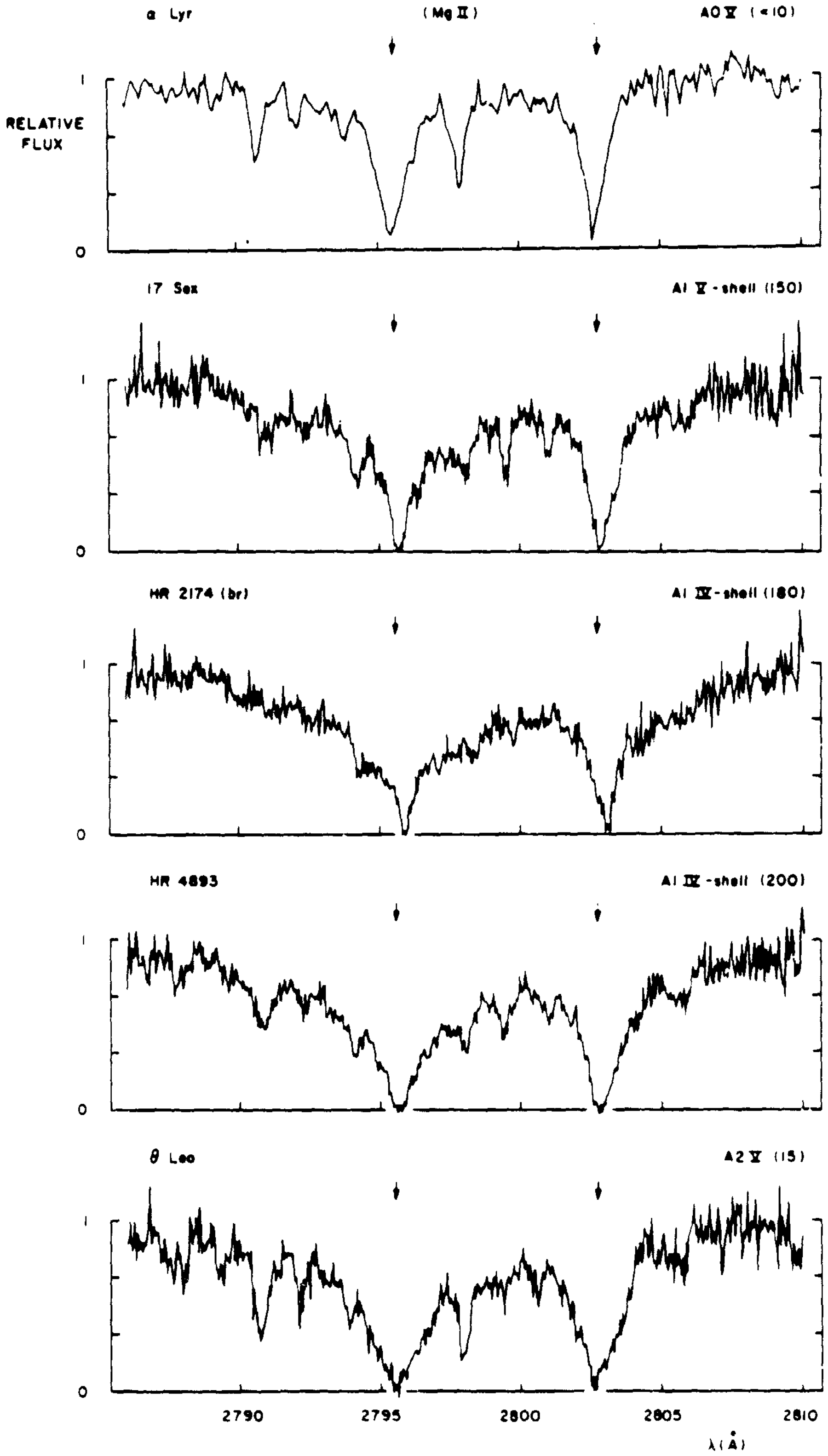

FIG. 17.—Spectral region including the Mg II resonance doublet at 2795.528 and 2802.704 Å (rest wavelengths indicated by arrows) in three A1 shell stars and an A0 V and A2 V standard star.

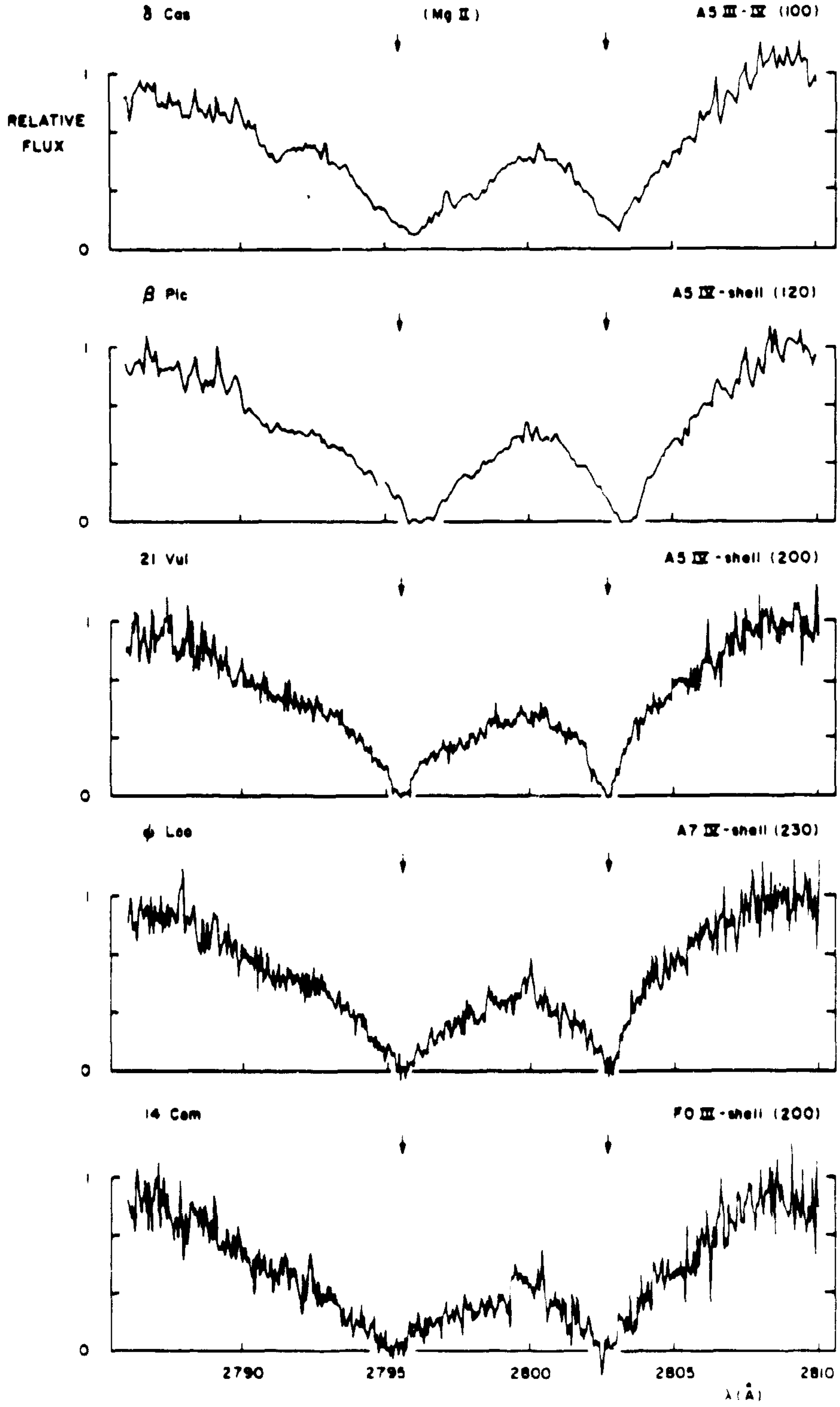

FIG. 18.—Spectral region including the Mg II resonance doublet at 2795.528 and 2802.704 Å (rest wavelengths indicated by arrows) in four A5–F0 shell stars and an A5 III–IV standard star.

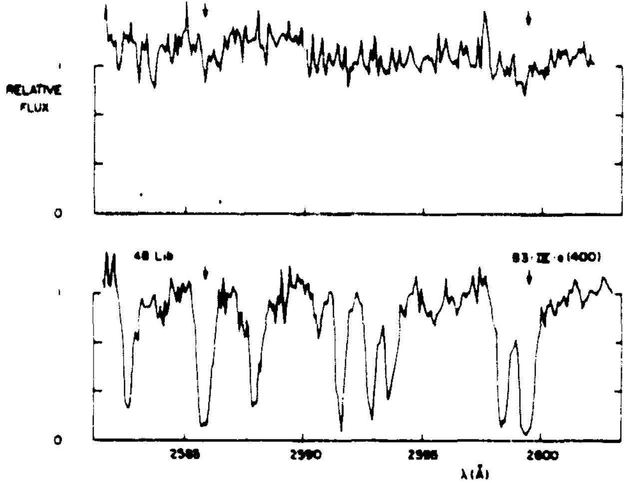

FIG. 19.—Spectral region including the strongest lines of the Fe II multiplet UV 1 in the B3 shell star 48 Lib with the spectrum of the B3 V standard star η UMa shown for comparison. The Fe II lines at 2585.876 and 2599.395 Å are indicated by arrows.

show ultraviolet emission lines, as may be seen in the above figures and, for the Mg II resonance lines, in Figures 17 and 18.

As might be expected, the Fe II lines which arise either from ground states (e.g., multiplet UV 1) or from metastable levels (see Figs. 22 and 23) are stronger, deeper, and sharper in the A–F type shell stars than in the standard stars. At least one star (β Pic) shows structure in the Fe II line profiles, suggesting velocity differences in the shell. Figures 17 and 18 also show that the Mg II resonance lines (analogously to the Ca II H and K lines in the optical region of the spectrum) are stronger and deeper in the A–F type shell stars than in the standard stars.

VII. SUMMARY AND CONCLUSIONS

We may summarize our results as follows:

1. Superionization in the Be stars extends to the latest spectral subtypes (B8–B9) but does not seem to be present in the A–F type shell stars.
2. The superionized lines in the Be stars appear to be correlated with $v \sin i$ in the sense that they tend to be stronger in stars with large $v \sin i$ than in stars with small $v \sin i$ at a given spectral type. They are strongest in the Be shell stars, which are also characterized by large rotational velocities. These results suggest that the hot component of the circumstellar envelope from which the superposed lines originate is not distributed with spherical symmetry, as has been suggested earlier by Marlborough and Peters (1983).
3. Superionization is also observed in our sample in normal, non-emission-line stars of spectral type as late as B5 and, possibly, as late as B7–B8, but not later. There is no obvious correlation in the strength of the superionized lines with $v \sin i$.
4. Asymmetrical or violet-displaced resonance lines, suggesting mass loss, were observed in all of our Be program stars but one, and also in a number of our standard stars, but not in the A–F type shell stars.
5. Computed lower limits to the mass-loss rates for the Be stars are in the range 1.4×10^{-12}–2.2×10^{-10} $M_\odot$ yr^{-1}. The results from the Si IV λ1393.755 line range between 5.3×10^{-12} and 3.5×10^{-11} $M_\odot$ yr^{-1}. These numbers are consistent with results for Be stars in common with earlier investigators.
6. The five Be shell stars as a group show the largest lower limits to the mass-loss rates of the stars in our sample.
7. The terminal velocities measured from the asymmetrical or violet-displaced resonance-line profiles range between 200 and 740 km s^{-1}, but show no clear correlation with ionization potential. This result is inconsistent with the usual assumption that ionization increases outward.
8. Mass loss and luminosity are correlated for the stars in our sample in the sense that the B-type stars show mass-loss effects while the (lower luminosity) A-type stars do not. Among the B-type stars themselves, the more luminous objects statistically have larger lower limits to the mass-loss rates than the less luminous stars.
9. Mass loss and effective temperature are correlated for the stars in our sample in the sense that the B-type stars show mass-loss effects while the (cooler) A-type stars do not. No correlation is found among the B-type stars themselves.
10. Mass loss and rotation appear to be uncorrelated for the stars in our sample.

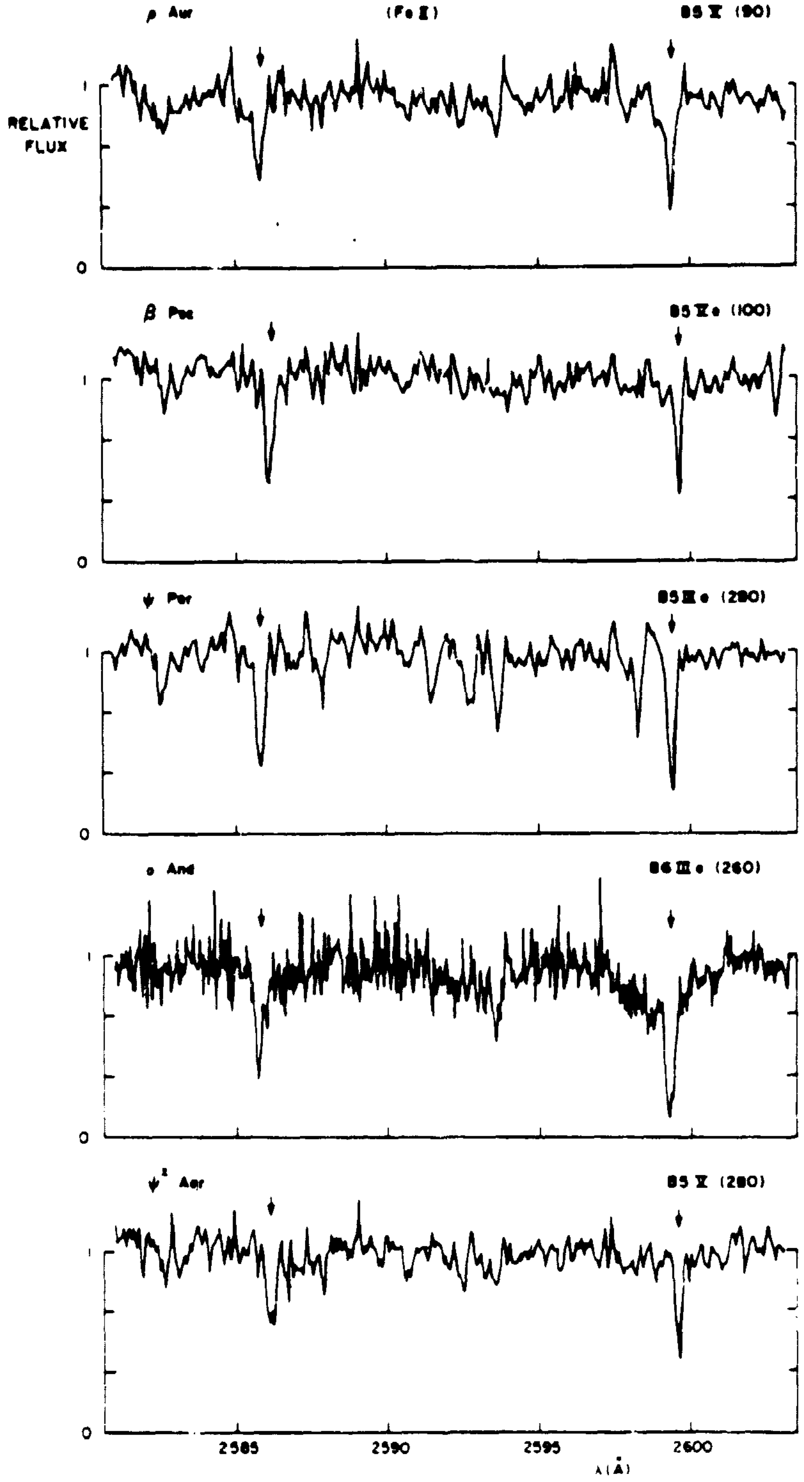

FIG. 20.—Spectral region including the strongest lines of the Fe II multiplet UV 1 in three B5–B6e stars and two B5 V standard stars. The Fe II lines at 2585.876 and 2599.395 Å are indicated by arrows.

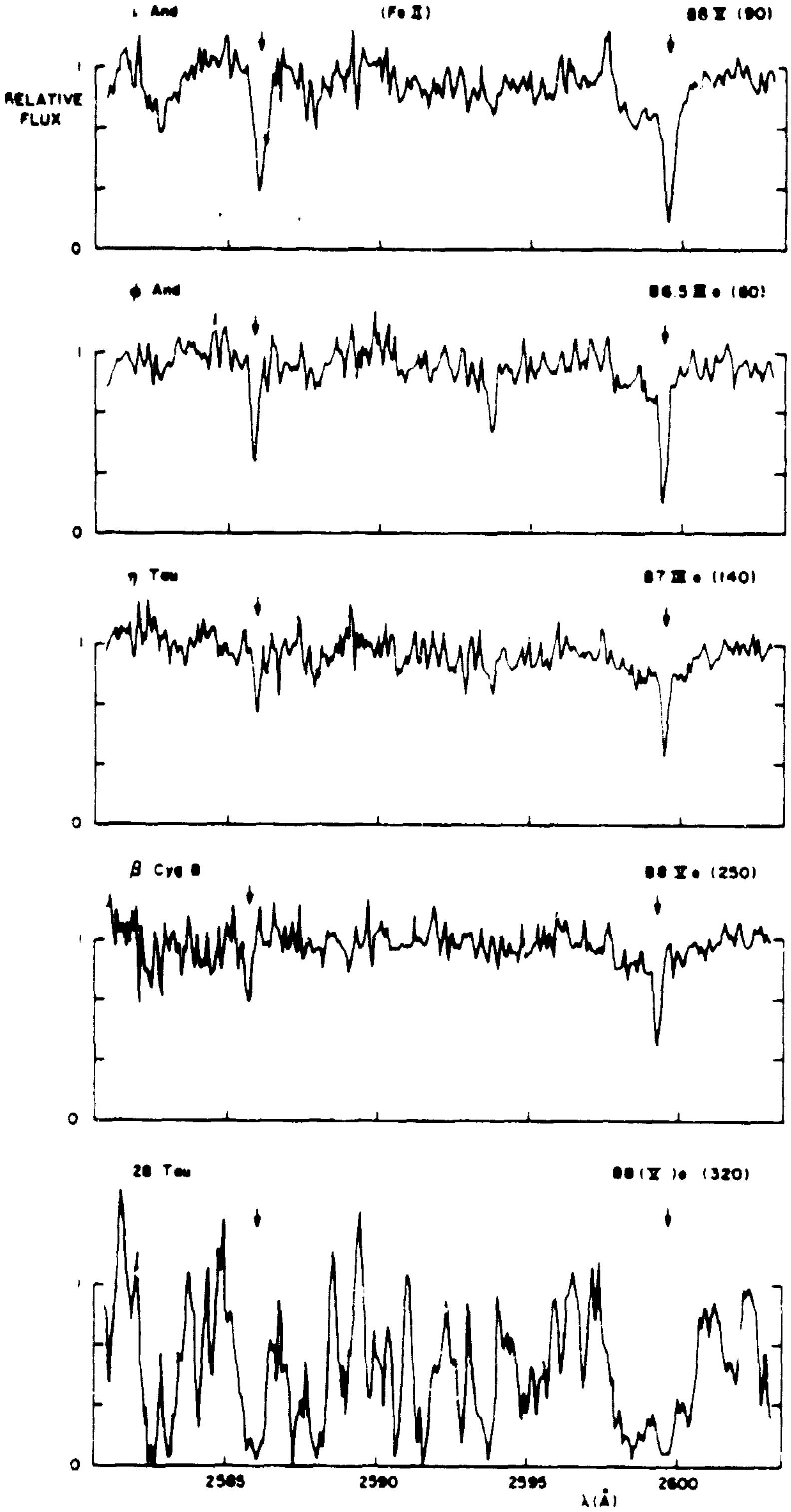

FIG. 21.—Spectral region including the strongest lines of the Fe II multiplet UV 1 in four B6.5–B8e stars and a B8 V standard star. The Fe II lines at 2585.876 and 2599.395 Å are indicated by arrows.

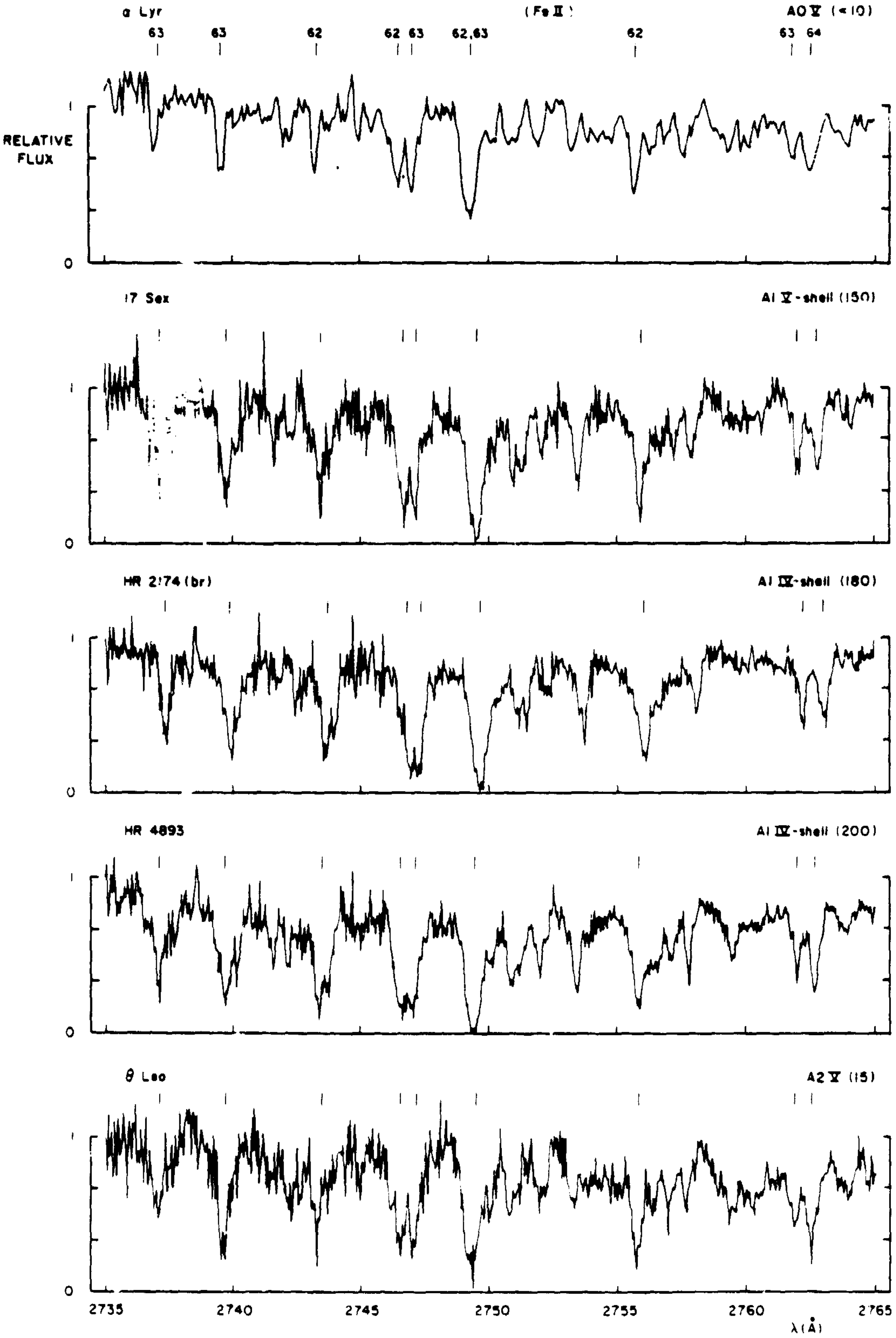

FIG. 22.—Spectral region including the strongest lines of the Fe II multiplets UV 62, UV 63, and UV 64 in three A1 shell stars and an A0 V and A2 V standard star.

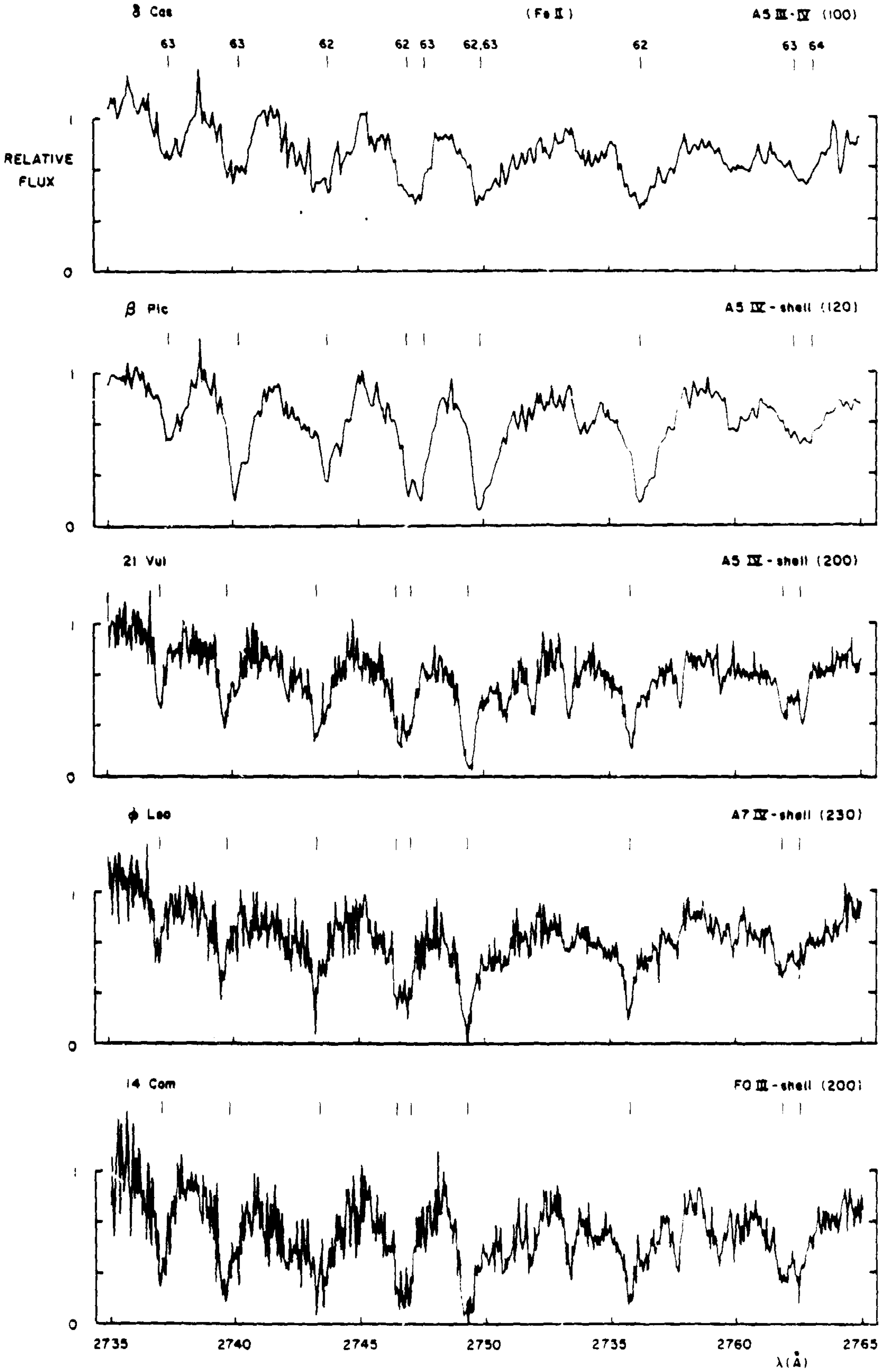

FIG. 23.—Spectral region including the strongest lines of the Fe II multiplets UV 62, UV 63, and UV 64 in four A5–F0 shell stars and an A5 III–IV standard star.

11. The only definite emission we observe is in the wings of the Mg II resonance doublet in two B5e stars in our sample—no Be stars of later type nor any of the A–F type shell stars show definite emission in any ultraviolet line.

12. All of our A–F type shell stars show strong ultraviolet Fe II and Mg II absorption spectra relative to standard stars of similar spectral type.

From these results, we draw the following conclusions:

i) Although Be stars are indeed nearly indistinguishable from normal B stars in the ultraviolet, as has been pointed out by Snow (1982*c*) and Doazan (1982), there are differences in degree: Be stars show superionization to later spectral types, generally have stronger superionized lines at a given spectral type, and, at least for the shell stars, show larger mass-loss rates.

ii) If correlations between the existence of superionized lines and mass loss versus $v \sin i$ exist for Be stars, they are weaker than correlations (from optical spectra and polarization measurements) involving the cool circumstellar envelope. This suggests that the latter is more nearly equatorially confined than the hot envelope which gives rise to the superionized lines and mass loss.

We are grateful to Mike Marlborough, Ted Snow, Paul Barker, and George Collins for very useful discussions. Some of this work was done while one of us (A. S.) was a visitor at the Institute for Astronomy of the University of Vienna, and the kind hospitality offered there is gratefully acknowledged. We thank also the *IUE* team at Goddard Space Flight Center for their help and cooperation, and the Ohio State University Instruction and Research Computer Center for providing computer time and facilities. The many drawings in this paper were prepared by Peter D. Stoycheff. We are also grateful to Mark Wagner for obtaining optical spectrograms of 17 Sex for us. Finally, we thank NASA for support via grant NAG 5-52.

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ROTATIONAL VELOCITIES OF LATER B TYPE AND A TYPE STARS AS DETERMINED FROM ULTRAVIOLET VERSUS VISUAL LINE PROFILES

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ABSTRACT

Theoretical line profiles of the Si III 1299 and Fe II 2756 lines are computed for rotating B5–A7 model stars, and compared with observed profiles from *IUE* spectra to derive rotational velocities. Real differences in widths for ultraviolet as compared with visual line profiles exist in our sample of B type stars (but not for the A type rapidly rotating stars), although these are not as large as previously reported in the literature. Comparison with our theoretical line profiles gives rotational velocities that are in good agreement with visually determined $v \sin i$ values for the same stars, which suggests that our shape-distorted, gravity-darkened models are reasonable.

Subject headings: line profiles — stars: early-type — stars: rotation — ultraviolet: spectra

I. INTRODUCTION

Most recently, Marlborough (1982) reviewed and commented on the observed discrepancy between rotational velocities ($v \sin i$) of Be stars as estimated from widths of absorption lines in ultraviolet as compared with visual spectra. He pointed out that perhaps the first to find such an effect were Morton *et al.* (1972), who observed several ultraviolet absorption lines (primarily C II and C III) in the rapidly rotating star ζ Ophiuchi to be unusually narrow. They suggested that the narrow lines may be formed predominantly in the more slowly rotating polar regions or that they may originate in a shell that has a lower rotational velocity than the visual photosphere. Not long after, Heap (1975, 1976, 1977) found ultraviolet lines of C III, C IV, Si III, and Si IV in the spectrum of the shell star ζ Tauri to be relatively narrow, corresponding to a considerably lower rotational velocity than is estimated from visual line profiles.

Hutchings (1976*a*), in attempting to interpret Heap's (1975, 1976, 1977) observations, suggested also that in the far-ultraviolet spectrum of a star with a temperature gradient across its surface, the underlying continuum radiation may come preferentially from polar regions, where $v \sin i$ is expected to be small. He followed this up (Hutchings 1976*b*) with a quantitative analysis, finding that $v \sin i$ from an assumed line profile at 4500 Å may be twice that at 1000–1200 Å under certain conditions. Sonneborn and Collins (1977) performed a similar but more refined analysis by assuming a rotationally distorted star and including the temperature and gravity dependence of actual lines. Their results were qualitatively similar to those of Hutchings (1976*a*, *b*), but they found a smaller variation in rotational line broadening than that predicted by Hutchings or observed in ζ Tauri by Heap.

There are problems in both the observational and theoretical areas, however, as Marlborough (1982) pointed out. In light of more recent ultraviolet observations of a number of Be stars, Heap's (1975, 1976, 1977) assumptions that all strong resonance lines in the ultraviolet are photospheric and that the circumstellar envelope would not contribute to lines of high stages of ionization are probably not correct. On the theoretical side, not only rotational shape distortion and the temperature-gravity variations of the line profile across the gravity-darkened stellar disk should be included but also (in principle) the effects of stellar winds, nonradial pulsations, and (possibly) differential rotation. All of the latter may affect the line profiles, but cannot easily be taken into account at this time.

As Hutchings (1976*b*) and Sonneborn and Collins (1977) have pointed out, if rotational broadening is indeed a measurable function of wavelength, it becomes possible in principle to gain knowledge of global properties of rotating stars and, specifically, to determine v and i separately. Indeed, such determinations have been attempted in recent years by Hutchings and Stoeckley (1977), Hutchings, Nemec, and Cassidy (1979), and Ruusalepp (1982). Are these results meaningful, in view of the uncertain assumptions and difficulties that have been pointed out by Marlborough (1982)?

In the hope of clarifying the situation somewhat, we have calculated new theoretical line profiles based on line-blanketed rotating model atmospheres, which include shape distortion and gravity darkening, plus the latest atomic data, and compared our results with ultraviolet line profiles in a number of later B type and A type stars observed with the *IUE* satellite.

II. THE THEORETICAL ULTRAVIOLET LINE PROFILES

Atomic absorption lines chosen for rotational velocity analysis should be (1) as free as possible from broadening or profile-distorting mechanisms other than rotation and (2) sufficiently strong and isolated from other lines to allow accurate profiles to be determined. These conditions present difficulties in the ultraviolet, where the lines tend to be crowded together and the strongest lines are inevitably resonance lines. Resonance line profiles can indeed often be measured with high precision but are also likely to be distorted by stellar winds, in normal B type as well as in Be stars, even in the later B types (cf. Marlborough 1982; Snow 1982; Slettebak and Carpenter 1983).

We have chosen two moderately strong nonresonance lines, which we assume to be photospheric and broadened essentially by rotation in our sample of later B type and A type stars. For the spectral range B5–B8, we use the feature near 1299 Å which is produced by the Si III (UV4) lines at 1298.960 and 1298.891 Å. These lines have their lower states at about 6.5 eV and their

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[1] Guest Investigator, *International Ultraviolet Explorer* satellite.

higher energy levels near 16.1 eV. The neighboring Si III lines at 1297 and 1301 Å were included in our calculations but were judged to be too weak in the observed spectra to be useful in the analysis. Finding suitable lines for the A type stars is more difficult, since they have little energy shortward of 1500 Å and the spectra are crowded with lines. The best compromise we could make was the 2755.733 Å line of multiplet UV62 in the spectrum of Fe II. This line is fairly strong and relatively isolated, and is produced from a lower energy level at 0.98 eV to a higher level at 5.46 eV.

While not resonance lines, both the Si III λ1299 and Fe II λ2756 lines have lower levels that do not connect strongly to the ground state. In this sense, these are metastable levels, and absorption lines arising from them will be enhanced under low-density conditions, which will overpopulate these levels. Our lines, therefore, should not be used for determining rotational velocities of shell stars, since they would yield estimates too low relative to the values of $v \sin i$ measured from lines that originate solely in the photosphere.

Synthetic spectra of regions 6 Å wide, approximately centered on each of the two lines, were computed by using a modified version of the ATLAS6 (Kurucz 1979) model atmosphere code, the global intensity integration routines of Collins and Harrington (1966), and previously computed model atmospheres described in Slettebak, Kuzma, and Collins (1980). The physical parameters for these line-blanketed model atmospheres, which account for the shape distortion and gravity darkening caused by rapid rotation (resulting in the variation of the photospheric continuum brightness with latitude), are given in the latter reference. Data for the atomic lines included in each of these synthetic spectra were obtained by screening the million-line list of Kurucz (1979) to obtain data for the strongest lines in each wavelength region under the physical conditions expected in each model. Since data for the Si III (UV4) lines are not in the Kurucz (1979) compilation, their wavelengths, *gf*-values, and lower energy levels were taken from Wiese, Smith, and Miles (1969). The Kurucz data were screened by computing approximate line-center opacities corresponding to the temperatures, pressures, and densities near Rosseland optical depth $\tau \sim \frac{2}{3}$ in the midlatitude atmospheres ($\sim 45°$) of each stellar model. The 100–200 strongest lines were segregated by retaining for each model only those lines with central opacities greater than 9.0 cm^2 g^{-1}. The Fe II line and all the background lines were computed using the wavelengths, energy levels, and *gf*-values from the Kurucz tape, and the broadening approximations described by Kurucz (1979) and Kurucz and Furenlid (1979). The approximate damping constants are defined as follows:

$$\Gamma(\text{Stark}) \approx 1 \times 10^{-8} N_e n_{\text{eff}}^5 ,$$

$$\Gamma(\text{radiative}) = \frac{2.223 \times 10^{15}}{\lambda^2} \quad (\lambda \text{ in Å}) ,$$

$$\Gamma(\text{van der Waals}) = 2.8393081 \times 10^{-10} (N_{\text{H I}} + 0.42 N_{\text{He I}}) T^{0.3} R^2 ,$$

where

$$R^2 = 2.5 n_{\text{eff}}^4 Z_{\text{eff}}^{-2}$$

and

$$n_{\text{eff}}^2 = \frac{13.595 Z_{\text{eff}}^2}{(\psi - E_{\text{lower level}} + 0.1)} ,$$

with $N_{\text{H I}}$ and $N_{\text{He I}}$ the number densities of H I and He I, respectively, n_{eff} the effective quantum number, and Z_{eff} the ionization stage. The ratio A of the damping to the Doppler width is defined by

$$A = \frac{\Gamma(\text{radiative}) + \Gamma(\text{van der Waals}) + \Gamma(\text{Stark})}{4\pi \Delta\nu_D} .$$

Collisional broadening of the Si III (UV4) lines was approximated using the semiclassical results of Sahal-Bréchot and Segre (1971), for which

$$\Gamma_e = Q_1 N_e T^{Q_2} .$$

The values of $Q_1 = 3.5 \times 10^{-5}$ and $Q_2 = -0.3$ used in this equation were taken from Kamp (1976). The radiation damping constants, Γ(radiative), for Si III (UV4) were approximated by the upper-state Einstein A-values. A microturbulent velocity of 2 km s^{-1} was assumed for all lines in all of the models. These spectra were calculated with a wavelength spacing $\Delta\lambda = 0.05$ Å, with the central wavelength of each line rounded to the nearest grid point to ensure that each line is represented in the spectrum.

Since rotational broadening dominates the profiles of the lines being considered for the $v \sin i$ values of interest, our results should not be strongly dependent on the above assumptions.

Table 1 lists the rotational parameters of the models for which line profiles were computed, including values of the fractional angular velocity $w = \omega/\omega_c$ (where ω_c is the critical angular velocity for which the centrifugal force at the star's equator balances the gravitational force), the equatorial rotational velocity v, and the inclination angle i between the rotation axis and the line of sight. Si III line profiles were calculated for the B5–A0 models, and Fe II line profiles for the A0–A7 models. Figures 1–3 show selected examples of these theoretical line profiles. For each profile, the full width of the line at half-maximum depth (FWHM) was measured, with the continuum taken to be at a residual intensity of 1.00, and assumed to be an indicator of the rotational velocity of that model. The signal-to-noise ratio in the observed *IUE* line profiles does not justify the use of additional line-width parameters in the analysis, in our opinion. Calibration curves with the measured FWHMs as a function of $v \sin i$ were then plotted for each line and model. Figure 4 shows such calibration curves, as an example, for the Fe II λ2756 line in A5 models.

TABLE 1

ROTATIONAL PARAMETERS OF MODELS FOR WHICH LINE PROFILES WERE COMPUTED

		$v \sin i$ (km s^{-1}) AT EACH SPECTRAL TYPE					
w	i	B5	B7	A0	A2	A5	A7
0	...	0	0	0	0	0	0
0.5	0°	0		0	0	0	0
	30	70		64	66	65	65
	45	98		91	93	92	92
	60	120		111	114	112	113
	90	139		128	132	129	130
0.9	0	0	0	0	0	0	0
	30	145	136	134	138	135	135
	45	205	193	189	195	191	191
	60	251	236	232	239	234	234
	90	290	273	268	276	270	270

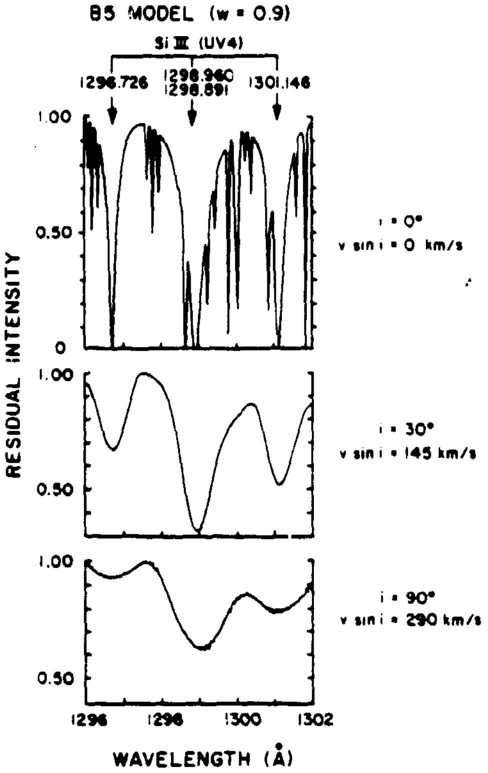

FIG. 1.—Theoretical line profiles of Si III multiplet UV4 lines in the wavelength range 1296–1302 Å, computed for B5 models with $w = 0.9$ and three values of the inclination i.

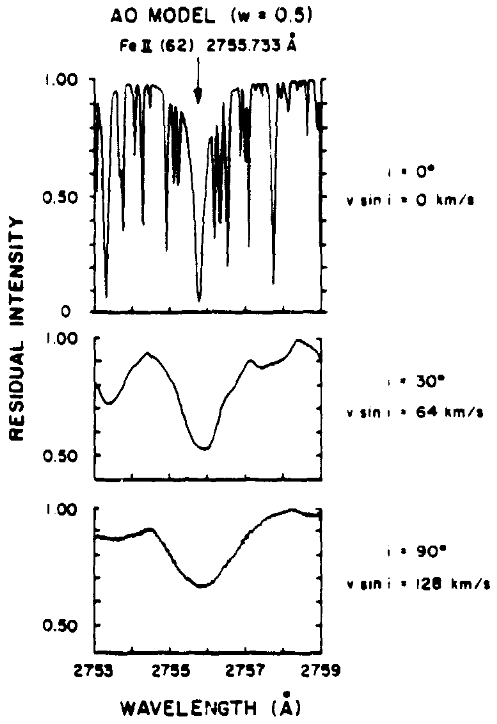

FIG. 2.—Theoretical line profiles of the Fe II multiplet UV62 line at 2755.733 Å, computed for A0 models with $w = 0.5$ and three values of the inclination i.

III. THE PROGRAM STARS

a) *Spectral Types and Rotational Velocities from Visual Spectra*

Since we already had *IUE* high-resolution spectra of a number of rapidly rotating standard stars of late B type and A type, we decided to concentrate on this spectral-type range. Our requirements were that the program stars have well-determined spectral types and rotational velocities (from visual spectra) and that high-resolution *IUE* spectra exist. We also tried to include stars with a range of rotational velocities in our sample, although highest priority was given to rapid rotators, where gravity-darkening effects are predicted to be the largest. Our 17 program stars are listed in Table 2. The spectral types and rotational velocities are from Slettebak *et al.* (1975), Slettebak (1982), and Slettebak and Carpenter (1983), or references therein. Details regarding how the visual $v \sin i$ values were determined, as well as error discussions, may be found in the aforementioned references. We estimate uncertainties of 10–15% of the values listed in Table 2.

b) *The Observed Ultraviolet Line Profiles and Rotational Velocities*

Table 3 lists the ultraviolet spectroscopic information for our program stars. All were observed with the *IUE* satellite in the high-dispersion mode, through the large aperture. We obtained eight of the spectra directly during the fifth and sixth observing episodes of the *IUE* (1982–1983) and the remaining nine from the *IUE* archives of the National Space Science Data Center (NSSDC). The star name and either the short-wavelength prime (SWP) or the long-wavelength redundant (LWR) camera image number are given in the first three columns of the table. Line widths (FWHM) were measured on the *IUE* spectra, corrected for the finite resolution of the *IUE* high-resolution camera, and compared with the line widths of the appropriate theoretical line profiles to obtain rotational velocities. Here an ambiguity enters, which is shown in Figure 4 and was discussed earlier by Collins (1974) and by Slettebak *et al.* (1975). These, as well as our present computations, show

TABLE 2

THE PROGRAM STARS

Star	HD	Spectral Type	Visual $v \sin i$ (km s^{-1})
ρ Aur	34759	B5 V	90
ψ^2 Aqr	219688	B5 V	280
β Psc	217891	B5 Ve	100
θ CrB	138749	B6 III(e)	320
φ And	6811	B6.5 IIIe	80
α Leo	87901	B7 V	280
η Tau	23630	B7 IIIe	140
β Cyg B	183914	B8 Ve	250
ι And	222173	B8 V	90
α Peg	218045	B9.5 III	120
δ Cyg	186882	B9.5 III	140
β Car	80007	A1 IV	125
α PsA	216956	A3 V	85
80 UMa	116842	A5 V	210
δ Cas	8358	A5 III–IV	100
α Oph	159561	A5 III	215
α Aql	187642	A7 V	225

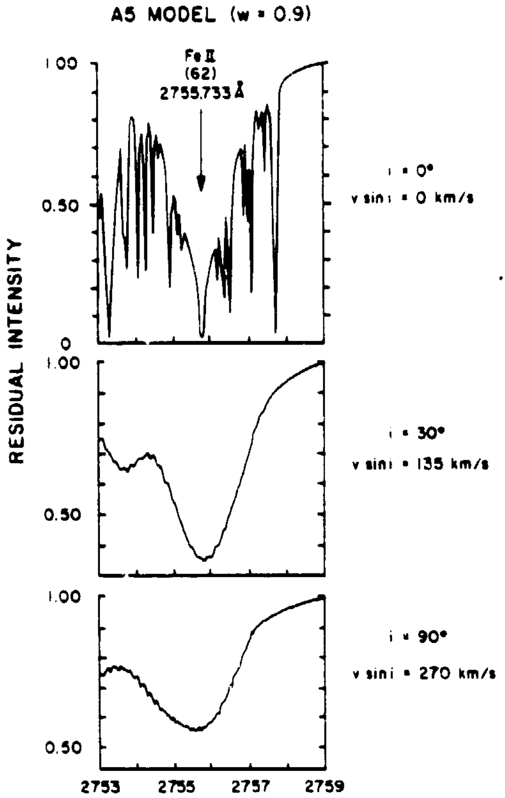

FIG. 3.—Theoretical line profiles of the Fe II multiplet UV62 line at 2755.733 Å, computed for A5 models with $w = 0.9$ and three values of the inclination i.

that the half-intensity width does not uniquely specify a rotational velocity for large values of w. We therefore took an average $v \sin i$ where both $w = 0.5$ and $w = 0.9$ models were applicable. This ambiguity introduces an uncertainty of 15–20% in our rotational velocities, listed in Table 3.

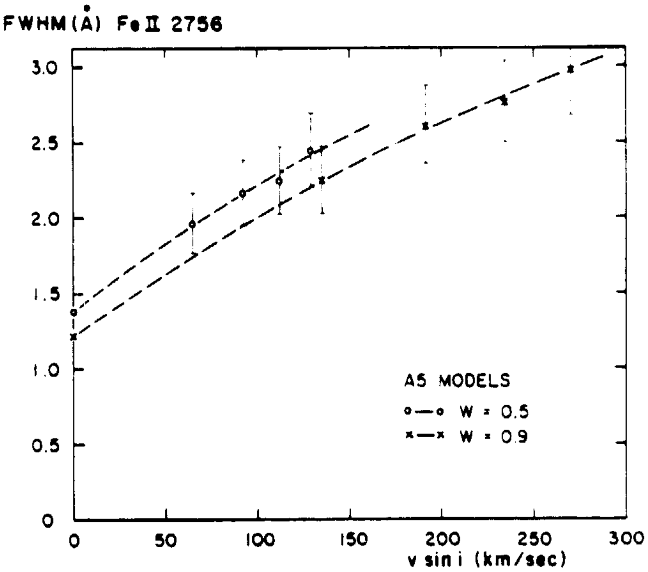

FIG. 4.—Line widths (FWHM) of theoretical line profiles of the Fe II λ2756 line in A5 model rotating stars, plotted against their rotational velocities, $v \sin i$. The circles and crosses represent computed values for $i = 0°$, 30°, 45°, 60°, and 90°; the dashed curves approximate the mean relations. Error bars of ±10%, which should represent maximum errors in view of uncertain continuum placements, are shown on the FWHM.

IV. RESULTS AND DISCUSSION

Figures 5 and 6 show the ultraviolet rotational velocities derived in this study from the Si III λ1299 and Fe II λ2756 lines, respectively, plotted against the visually determined $v \sin i$ values for the B5–B8 and B8–A7 program stars. The agreement between the values of $v \sin i$ is quite good in both cases, and we conclude that our models reasonably represent the observed line widths.

Sonneborn and Collins (1977) introduced the dimensionless parameter

$$Q = (\mathrm{FWHM}/\lambda)_1/(\mathrm{FWHM}/\lambda)_2$$

TABLE 3

ULTRAVIOLET DATA FOR THE PROGRAM STARS

STAR	CAMERA IMAGE NO. SWP	CAMERA IMAGE NO. LWR	Si III λ1299 FWHM (Å)	Fe II λ2756 FWHM (Å)	UV $v \sin i$ (km s^{-1})
ρ Aur	10389	...	0.80	...	70
ψ^2 Aqr	10385	...	1.52	...	255
β Psc	10386	...	0.86	...	85
	15512	...	0.84	...	
θ CrB	14431	...	1.63	...	335
	16288	...	1.69	...	
φ And	10387	...	0.77	...	65
χ Leo	10379	...	1.55	...	280
η Tau	10378	...	1.19	...	185
β Cyg B	5369	...	1.41	...	235
ι And	10376	...	0.82	...	85
		9057	...	1.31	80
χ Peg	...	2386	...	2.23	145
δ Cyg	...	3041	...	2.11	140
β Car	...	2970	...	1.69	105
χ PsA	...	11022	...	1.81	110
80 UMa	...	11712	...	2.57	190
δ Cas	...	11214	...	1.92	75
χ Oph	...	5925	...	2.72	220
χ Aql	...	2449	...	3.06	190
	...	3012	...	2.79	

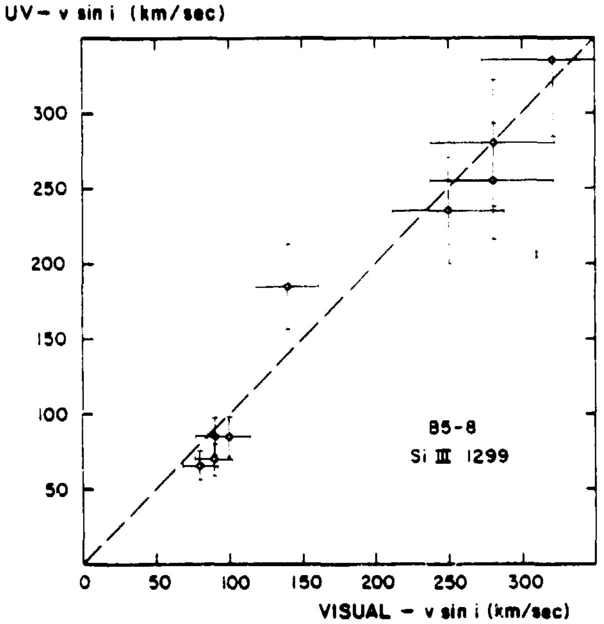

FIG. 5.—Visually determined rotational velocities (v sin i) plotted against ultraviolet v sin i values (from Si III λ1299 line profiles) for the B5–B8 program stars. Error bars of ±15% have been placed on each point. The dashed line is a 45° line representing perfect agreement between the visual and ultraviolet v sin i values.

as a measure of the difference in rotational line broadening as a function of wavelength. We can compute Q from our ultraviolet computations in this paper and the visual line profiles calculated by Slettebak, Kuzma, and Collins (1980). This gives values of Q between 1.2 and 1.3 for B5 models with $w = 0.9$ and $i = 90°$, and agrees quite well with observed Q-values for the rapidly rotating B5–B7 stars in this study obtained from *IUE* spectra compared with He I λ4471 line widths for the same stars in Slettebak *et al.* (1975). This confirms that there is an effect of wavelength dependence on rotational line broadening in the B type stars, that it can be understood in terms of gravity-darkened models, but that it is probably not as large as previously thought.

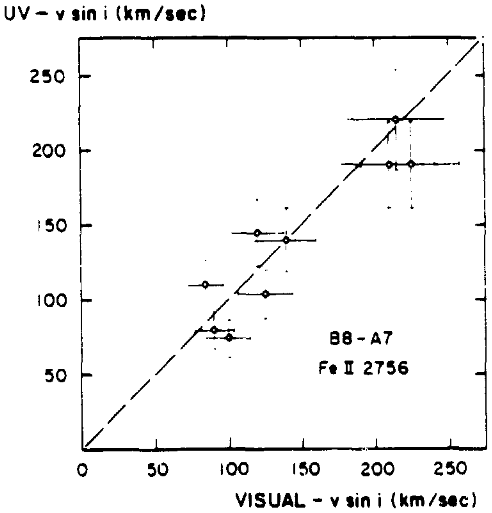

FIG. 6.—Visually determined rotational velocities (v sin i) plotted against ultraviolet v sin i values (from Fe II λ2756 line profiles) for the B8–A7 program stars. Error bars of ±15% have been placed on each point. The dashed line is a 45° line representing perfect agreement between the visual and ultraviolet v sin i values.

Our Q-values for the A type stars in this study, on the other hand, are near unity, even for the most rapidly rotating objects. There are probably several reasons for this. (1) Even the most rapidly rotating A type stars appear to have fractional angular velocities that are not sufficiently close to unity to produce striking gravity-darkening effects. (2) Since the temperature difference between pole and equator in a rapidly rotating A type star is considerably less than in a rapidly rotating B type star, the gravity-darkening effects would be expected to be much smaller in any case. (3) The Fe II λ2756 line, which we chose for analysis in the A type spectra, is not so widely separated in wavelength from the Mg II λ4481 line usually chosen for visual v sin i determinations as to show a large wavelength-dependence effect.

Before our conclusions are taken too seriously by the reader, however, we feel it is necessary to comment on the accuracy of our observed and theoretical line profiles, as well as on some ambiguities in our analysis:

1. Placing the continuum for the observed line profiles is extremely uncertain. The high-resolution *IUE* spectra are obtained with an echelle spectrograph, and establishing continuity between echelle orders is difficult and not always successful. The problem of locating the continuum becomes worse when the line profiles to be measured are broad and other strong lines are in the vicinity. Also, for broad lines, a small error in continuum placement can result in a relatively large error in measured line width and, therefore, rotational velocity.

2. There is also some uncertainty in the location of the continuum for the theoretical line profiles, because of the restricted (by limitations on available computing time) wavelength extent of the spectra computations.

3. Figure 4 shows two problems introduced in using the FWHM versus v sin i calibration curves. The first, which we have already discussed, is that there is an essential ambiguity in using line widths to estimate rotational velocities for rapidly rotating stars, as is shown by the difference between the $w = 0.5$ and $w = 0.9$ curves. A second source of uncertainty arises from the flattening out of the FWHM versus v sin i curves for large values of v sin i. A small difference in the measured line width thus leads to a large change in v sin i.

The above discussion suggests that Figures 5 and 6 be regarded with some caution. The rather good correlation between the visually determined and ultraviolet v sin i values shown there is probably somewhat fortuitous. On the other hand, we feel that our analysis is sufficiently accurate to reveal striking differences, and these do not seem to be present.

Improvements in both theory and observation could be made in future work. As we have already stated, we feel that our theoretical treatment is sufficiently accurate to deal with this problem in view of the fact that rotational broadening dominates the line profiles. It would be advantageous, however, to compute synthetic spectra covering a larger wavelength range than was possible in this paper, to define the theoretical continuum more precisely. On the observational side, we anticipate that the quality of the line profiles obtained with the high-resolution spectrograph on the Space Telescope would permit a comparison of the entire line profile, rather than only the FWHMs used in this investigation, and therefore would lead to more accurate results. The choice of suitable lines for analysis presents real problems, since all strong, unblended ultraviolet lines seem to arise either from ground states or from metastable levels. We have chosen what we consider to be the best lines for our purpose, but higher resolution, less noisy data may permit the use of more suitable lines.

The many uncertainties inherent in analyses involving large rotational line broadening also suggest to us that attempts to separate *v* and *i* from line profiles must be carried out with great caution, and, indeed, it is not clear to us that such a separation is even feasible.

We can summarize our results as follows. Real differences in widths for ultraviolet versus visual line profiles in rapidly rotating B type stars exist, but these are not as large as previously reported in the literature. This discrepancy may be due in part to the earlier use of resonance and metastable lines that have substantial nonphotospheric contributions from winds and/or circumstellar envelopes. The good agreement we obtain between ultraviolet and visually determined rotational velocities suggests that our shape-distorted, gravity-darkened models are reasonable and may bear some resemblance to the real world.

We acknowledge very helpful conversations with George Collins. The *IUE* team at Goddard Space Flight Center was most cooperative, as was Wayne Warren of the National Space Science Data Center at Goddard. We thank also the Ohio State University Instruction and Research Computer Center for providing computer time and facilities, and NASA for support via grant NAG 5-52.

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a) Ultraviolet spectra of some brighter Lambda Bootis stars

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d) ULTRAVIOLET SPECTRA OF LAMBDA BOOTIS STARS

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f) Stellar Atmospheres

g) Supplement Series

[1] Guest Investigator: International Ultraviolet Explorer Satellite

Summary. - Strengths of ultraviolet lines from IUE spectra of six bright λ Bootis or suspected λ Bootis stars are compared with line strengths in standard "normal" stars of similar spectral types. In five of the six stars (the sixth, ε Sgr, appears to be normal in the ultraviolet) the lines of carbon and nitrogen are stronger than in the standard stars, while the lines of magnesium, aluminum, and (to a lesser extent) iron are weaker. These abnormal line strengths, which seem not to be due to luminosity effects, might be interpreted either in terms of diffusion processes or by nucleosynthesis. The line profiles of ultraviolet resonance lines do not show the asymmetries associated with mass loss in the hotter stars, but the low mass-loss rates suggested in the diffusion theory are probably not excluded by the observations. On the other hand, the observed strength of ultraviolet carbon and nitrogen lines is suggestive of nucleosynthetic processes, and the possibility exists that the λ Boo stars are evolved objects in which nuclearly-processed CNO elements have somehow found their way to the surfaces.

Key Words: UV spectra -- Lambda Bootis stars -- peculiar A stars -- stellar atmospheres

I. Introduction

Morgan et al. (1943) described the spectrum of λ Bootis, the prototype of this class of stars, as follows: "The spectral type of λ Boo is near A0, as far as can be determined. The spectral lines, while not unusually broad, are very weak, so that the only features easily visible are a weak K-line and the Balmer series of hydrogen". Additional members were proposed by Andersen and Nordström (1977), Graham and Slettebak (1973), Parenago (1958), Slettebak (1952, 1954, 1975), and Slettebak et al. (1968), but it was obvious that the class was rather ill defined. Hauck and Slettebak (1983) therefore proposed the following spectroscopic definition: "λ Boo stars are A-F type stars with metallic lines which are too weak for their spectral types, when the latter are determined from the ratio of their K-line to Balmer-line strengths. They are distinguished from weak-lined Population II (horizontal branch) stars by the fact that they have normal space velocities and moderately large rotational velocities." Based on this definition, they listed nine stars which appear to be λ Boo stars. Meanwhile, Abt (1984), suggesting that "... an easy way to discover λ Bootis stars is to search for stars near A0 that have weak (Mg II) 4481", presented a list of 26 stars near A0 selected from the Bright Star Catalogue which have weak Mg II 4481, and which he provisionally refers to as λ Bootis stars.

Abundance analyses of some of the brighter λ Boo stars have been carried out by a number of investigators. Burbidge and Burbidge (1956), using curve-of-growth methods, found λ Boo to have only about 10 percent of the "normal" abundance of Fe and only about 3 percent of the "normal" abundance of Ca, with smaller underabundance ratios for 29 Cyg. Kodaira (1967) found "λ Boo and

π^1 Ori to have normal oxygen abundance, in contrast to the underabundance of magnesium by a factor of more than 5." For 29 Cyg, on the other hand, he found the magnesium abundance to be "rather normal". Baschek and Searle (1969) carried out a curve-of-growth abundance analysis of a number of objects which had been classified as λ Boo stars, and found that "only λ Boo, 29 Cyg, and π^1 Ori form a distinct group from the composition point of view. The composition characteristics of the λ Boo group are (1) a deficiency of about a factor 3 in the iron-group elements; (2) nearly normal oxygen abundance; (3) deficiencies by Mg and Ca that are at least as large as those of the iron-group elements and sometimes are larger."

Baschek and Searle (1969) agreed with Sargent's (1965) suggestion that the λ Boo stars are related to the peculiar A stars, and stated further that "... they form a separate type of Ap star, and their surface composition does not reflect that of the material form which they formed." More recently, Michard et al. (1983) have suggested that the λ Boo stars may be related to the A-type metallic-line (Am) stars. They propose that the diffusion process (Michaud 1970) plus a mass-loss rate of the order of 10^{-13} $M_{\odot}$/yr could lead to the λ Boo phenomenon.

An ultraviolet study of nine λ Boo or suspected λ Boo-type stars was carried out by Baschek et al. (1984) using IUE spectra. Of five stars listed by Hauck and Slettebak (1983) as λ Boo stars based on optical spectra, they found that four (λ Boo, π^1 Ori, HR 4881, and HD 105058) also show λ Boo characteristics in the ultraviolet, while the fifth (29 Cyg) shows only marginal agreement. A study of high-resolution IUE spectra of λ Boo itself led them to conclude that "the line strengths indicate that carbon (and perhaps also nitrogen) may be normal or even enhanced in λ Boo in contrast to the underabundance of the heavier elements."

The work of Baschek et al. (1984) was not known to me when I also obtained IUE spectra of some of the brighter λ Boo stars in October 1983 and January 1984. My goal was to compare high-resolution IUE spectra of λ Boo stars with spectra of standard "normal" stars of similar spectral type, to determine which chemical elements have stronger, weaker, or normal lines in the λ Boo stars. This paper presents these results, based on independent IUE observations. No attempt is made to analyze the results or derive abundances, since the data do not warrant a refined analysis, in my opinion. The lines of certain elements do appear to be strong and others weak, relative to their appearance in standard stars, however, and these will be pointed out.

II. Observations

Table 1a lists the six λ Boo or suspected λ Boo-type stars for which high-resolution IUE spectra were obtained, including the image numbers. Table 1b gives the same information for 14 standard stars with normal optical spectra, of spectral types B8 to A5. Spectral types and rotational velocities for the peculiar stars are referenced in the paper by Hauck and Slettebak (1983), except for 15 And (see the Note to Table 1a), while those for the standard stars are from Slettebak et al. (1975) or references therein.

Identifying representative spectrum lines of various chemical elements which are unblended and at the same time strong enough to permit accurate equivalent width measurements is a near impossibility for ultraviolet spectra of hot stars. Furthermore, many or most such candidate lines are resonance lines, which introduces complications due to interstellar and/or chromospheric contributions. Recognizing these difficulties, Table 2 lists the lines (many of them blended) which seemed least objectionable for the comparison of line strengths, which is the purpose of this investigation.

IUE calcomp tracings of the spectra of all the program stars and all the standard stars but two (18 Tau and λ UMa) were available with a paper scale of 2 Å/inch. Equivalent widths were measured on these tracings, using a planimeter. In addition, equivalent widths of certain lines in the spectra of 18 Tau and λ UMa, as well as in some of the program stars, were measured using a reduction program at the IUE Regional Data Analysis Facility in Boulder, Colorado. Because a considerable range of continuum is necessary on either side of each spectrum line to be measured in order to obtain meaningful equivalent widths, the Boulder reduction program was found, unfortunately, to be useful only for certain lines.

III. Results

Baschek et al. (1984) have pointed out the difficulties in reaching quantitative conclusions regarding the abundance of carbon in the atmosphere of λ Boo in view of saturation effects in the lines (making them sensitive to microturbulent velocities and damping constants) and the possibility of differential non-LTE effects on the line strengths. There are also a number of difficulties on the observational side: (1) As has already been mentioned, ultraviolet spectra of the hotter stars are crowded with lines and it is extremely difficult to find lines which are unblended and at the same time strong enough to permit accurate equivalent-width measurements; (2) the tracings utilized in this study often had poor flux corrections at the ends of the echelle orders ("ripple" corrections), which made continuum placement for lines at these wavelengths very difficult; (3) unfortunately, a number of my tracings were very noisy, which affected both the line profiles and continua.

For all of the above reasons, only line strengths of the program stars versus standard stars are presented in this paper, with no attempt at analysis. My hope is that, by measuring the lines in the λ Boo stars and in the standard stars of similar spectral types in the identical manner, meaningful comparisons can be made.

The results are shown in Figures 1-24, one for each of the spectrum lines listed in Tables 2a and 2b. In each case, equivalent width W(Å) is plotted against spectral type for both the program and the standard stars.

An inspection of Figures 1-24 shows, first, a considerable scatter in the points, making it difficult to decide whether a given line in the program stars is weaker, stronger, or comparable to the line in the standard stars. In a number of cases, the comparison shows a definite difference, however:

1. Figures 3 and, especially 7, suggest that the lines of carbon are decidedly stronger in the λ Boo stars (with the exception of ε Sgr) than in the standard stars of comparable spectral type. This supports the conclusion of Baschek et al. (1984) that carbon may be normal or even enhanced in λ Boo.
2. Figures 4, 5, 10, and 11 suggest that the lines of nitrogen are also decidedly stronger in the λ Boo stars (again, with the exception of ε Sgr) than

in the standard stars of comparable spectral type. Baschek et al. (1984) had also suggested that the lines of nitrogen in λ Boo "perhaps" may be normal or even enhanced. The N I (multiplet UV 4) lines at 1492.63, 1492.82, and 1494.68 Å are, unfortunately, blended with the C I (multiplet 64.06) 1494.74 Å and C I (multiplet 64.08) 1492.74 Å lines in an unknown way, but Figures 10 and 11 show that the N I (multiplet UV 9) lines at 1743 and 1745 Å are also strong relative to standard stars.

3. Figures 21-24 show that the lines of magnesium are decidedly weaker in the λ Boo stars (again, with the exception of ε Sgr) than in the standard stars of comparable spectral type. This might be expected since the weakness of Mg II 4481 is one of the criteria for identifying λ Boo stars. Burbidge and Burbidge (1956), Kodaira (1967), and Baschek and Searle (1969) all found magnesium to be underabundant in the λ Boo stars.

4. Figure 8 shows that the Aℓ II 1671 resonance line is considerably weaker in the λ Boo stars (again, with the exception of ε Sgr) than in the standard stars of comparable spectral type. It is interesting to note that Sadakane et al. (1983) found the resonance lines of both Aℓ II and Aℓ III to be abnormally weak in many B-type chemically peculiar stars.

5. Figures 14, 16, 17, 19, and 20 show the lines of Fe II to be definitely weaker or slightly weaker in the λ Boo stars (again, with the exception of ε Sgr) than in the standard stars of comparable spectral type. The effect is not as pronounced as for the Mg II and Aℓ II lines, but seems to be real.

6. Figure 9 shows the Ni II 1742 line to appear slightly weaker, on the average, in the λ Boo stars (again, with the exception of ε Sgr) than in the

standard stars of comparable spectral type. This effect may not be real, in view of the uncertainty of the equivalent width measurements.

7. No definite conclusions can be drawn about the strengths of the lines of silicon (Figure 6), sulfur (Figure 12), or manganese (Figures 15 and 18), in view of the uncertain measurements, except to suggest that the lines seem neither to be conspicuously strong nor conspicuously weak relative to the same lines in the standard stars.

A number of other conclusions may be drawn from the data:

The bright star ε Sgr seems not to be a λ Boo star on the basis of the strengths of its ultraviolet lines. In addition to the results already mentioned, Figures 1, 2, and 13 show that lines of Si III, O I, and Aℓ III, respectively, in ε Sgr appear to have normal strengths relative to standard stars of similar spectral type. On the other hand, the optical spectrum of ε Sgr is peculiar in that the Mg II 4481 line is broad and rather weak while the Balmer absorption cores, Fe II 4233, and the Ca II K-line are considerably narrower, as was pointed out earlier (Slettebak 1975). The spectrum could be composite, but again there is no evidence for abnormal line intensities in the ultraviolet. Hauck and Slettebak (1983) have pointed out that the colors of ε Sgr are unusual and that its position in both a d vs. (B2-V1) diagram and a c_1 vs. (b-y) diagram place it in the area of the horizontal-branch A-type stars. This object does indeed deserve further study.

Abt's (1984) suggestion that 15 And is a λ Boo star on the basis of its weak Mg II 4481 line seems to be confirmed by the IUE data. The star shares the strong C I and N I lines, the weak Mg II and Aℓ II lines, and the slightly weak Fe II lines shown by the other λ Boo stars studied in this paper. On the other hand, Hauck (1985) decided against membership of 15 And in the λ Boo group on the basis of its colors in the Geneva photometric system. Hauck's criteria for membership are severe, however, and he states that "some stars (of the objects investigated by him) could marginally be members of the group .... with only one clear exclusion (an obvious supergiant)".

In view of Michaud et al.'s (1983) suggestion that the λ Boo stars may be losing mass, the ultraviolet resonance line profiles were examined for asymmetries in all of the program stars. No asymmetries were found. On the other hand, Michaud et al.'s suggested mass-loss rate of 10^{-13} $M_{\odot}$/yr is not much larger than the mass loss due to the solar wind -- under these circumstances there would probably be no measurable effect on the line profiles.

IV. Conclusions

1. Five of the six λ Boo stars investigated in this study have similar ultraviolet spectral characteristics: the lines of carbon and nitrogen are stronger than in standard stars of similar spectral type, while those of magnesium, aluminum, and (to a lesser extent), iron are weaker. Nickel may also be weakened.
2. The sixth star investigated, ε Sgr, while showing peculiarities in its optical spectrum and colors, looks normal for its spectral type in the ultraviolet.
3. The data presented in Figures 1-24 show no striking differences in line strengths between main-sequence and giant stars, suggesting that the behavior

of the ultraviolet lines in the λ Boo stars is not due to luminosity effects. The abnormal line strengths in the ultraviolet spectra of the λ Boo stars might then be interpreted either in terms of diffusion processes, as suggested by Michaud et al. (1983), or by nucleosynthesis. The line profiles of ultraviolet resonance lines do not show the asymmetries associated with mass loss in the hotter stars, but the low mass-loss rates suggested by Michaud et al. are probably not excluded by the observations. On the other hand, the observed strength of ultraviolet carbon and nitrogen lines is suggestive of nucleosynthetic processes (Kodaira 1967 and Baschek and Searle 1969 had earlier found that the oxygen abundance is normal in λ Boo stars). The possibility thus exists that the λ Boo stars are evolved objects in which nuclearly-processed CNO elements have somehow found their way to the surfaces.

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Table 1a. IUE observations of λ Boo and suspected λ Boo-type stars

Star	HD	Sp. Type	v sin i (km/s)	SWP	LWP
π^1 Ori	31295	A0p	110	21318	2098 2557
HR 4881	111786	A3p	-	21945	2561
λ Boo	125162	A0p	100	21319 21320	2099
ε Sgr	169022	B9p	100	21316 21317	2097
29 Cyg	192640	A2p	90	21321	2100
15 And*	221756	A1p	100	21944	2559

* Note: Abt (1984), who first listed 15 And as a probably λ Boo star, gives A1Vp (Mg wk) as the spectral type with v sin i = 109 km/s. On a Perkins 72" spectrogram of Oct. 23, 1985 I confirm the spectral type to be near A1, the Mg II 4481 line does indeed appear weak for that type, and the rotational velocity is in the neighborhood of 100 km/s.

Table 1b. IUE observations of standard stars

Star	HD	Sp. Type	v sin i (km/s)	SWP	LWR
δ Cas	8358	A5 III-IV	100	11214	9831
β Ari	11636	A5 V	70	-	8712
18 Tau	23324	B8 V	250	7923	6901
γ Gem	47105	A0 IV	< 10	15542	12021
β Car	80007	A1 IV	125	3383	2970
λ UMa	89021	A2 IV	35	8187	7115
θ Leo	97633	A2 V	15	15505	11981
α Dra	123299	A0 III	15	3840	3423
109 Vir	130109	A0 V	330	-	1922
α Lyr	172167	A0 V	< 10	11213	9830
δ Cyg	186882	B9.5 III	140	14822	3041
α PsA	216956	A3 V	85	14417	11022
α Peg	218045	B9.5 III	120	2671	2386
ι And	222173	B8 V	90	10376	9057

Table 2a. Spectrum lines measured in the IUE short-wavelength range

Element	Multiplet	Wavelength (Å)
Si III	UV 4	1298.89, 1298.96
O I	UV 2	1302.17
C II	UV 1	1334.53, 1335.71
N I	UV 4	1492.63, 1492.82, 1494.68
Si II	UV 2	1533.44
C I	UV 2	1656.27, 1656.93, 1657.01, 1657.38, 1657.91, 1658.12
Al II	2	1670.81
Ni II	5	1741.56
N I	UV 9	1742.72, 1742.73 1745.25, 1745.26
S I	2	1820.37
Al III	1	1862.78
Fe II	126	1864.74

Table 2b. Spectrum lines measured in the IUE long-wavelength range

Element	Multiplet	Wavelength (Å)
Mn II	1	2576.11 2605.70
Fe II	1	2585.88 2598.37, 2599.39
Fe II	63	2739.54
Fe II	62, 63	2749.18 (63), 2749.32 (62), 2749.48 (63)
Mg II	1	2795.52 2802.70
Mg II	2	2928.62 2936.50

CAPTION FOR FIGURES

FIGURES 1-24. - Equivalent widths W(Å) of the ultraviolet lines versus spectral types of the program and standard stars. The λ Boo and suspected-λ Boo stars are plotted as squares, with symbols designating each star as follows: $\pi = \pi^1$ Ori, HR = HR 4881, λ = λ Boo, ε = ε Sgr, 29 = 29 Cyg, 15 = 15 And. A colon behind any symbol indicates that the equivalent width measurement is particularly uncertain. The dashed line in each figure has been drawn by eye through the standard-star points, as a guide to the general trend of the line strengths for the standard stars; it has no physical meaning beyond that.

Fig. 1

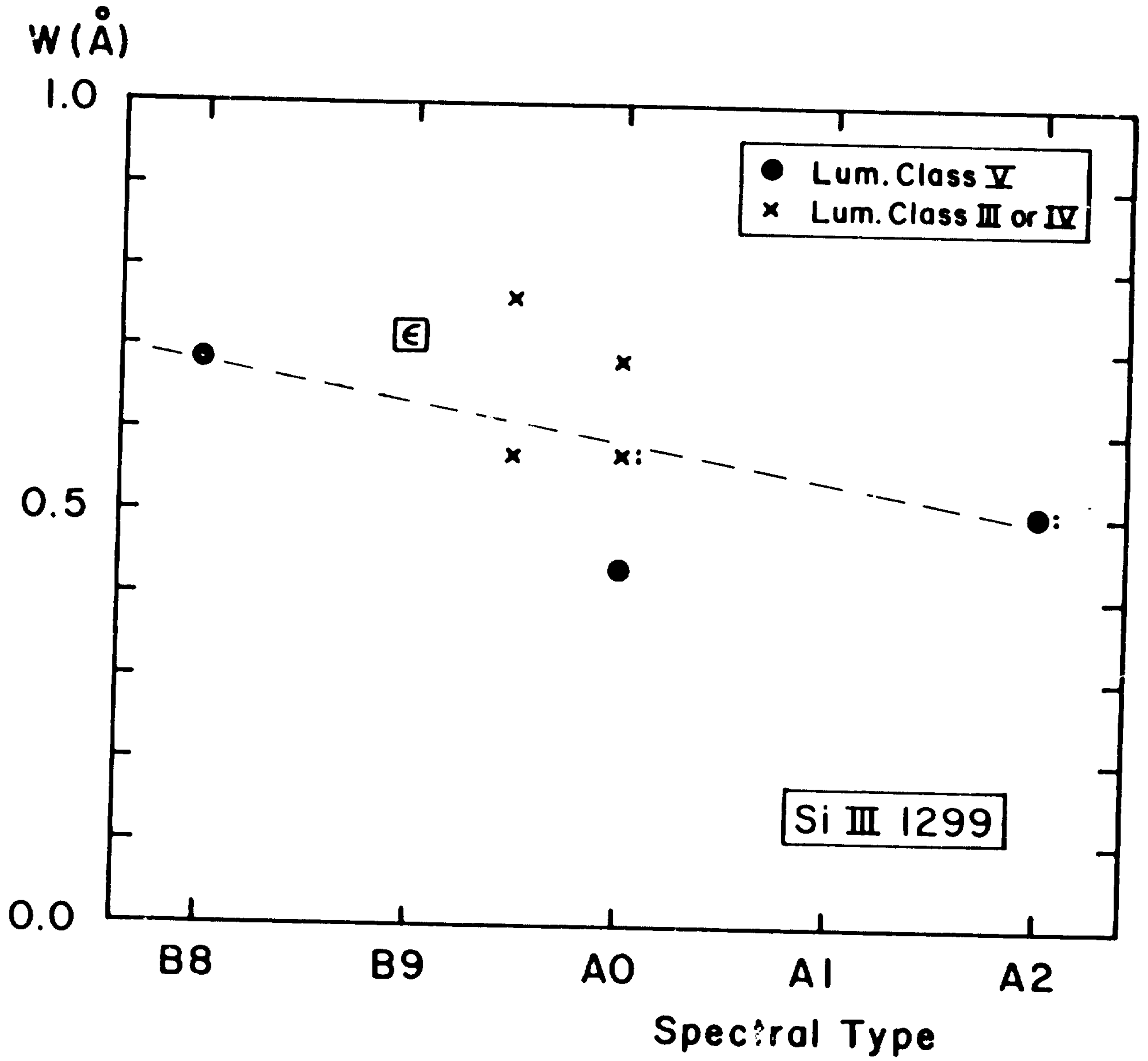

W (Å)
1.0
0.5
0.0
● Lum. Class V
× Lum. Class III or IV
ϵ
Si III 1299
B8
B9
A0
A1
A2
Spectral Type

FIG. 2

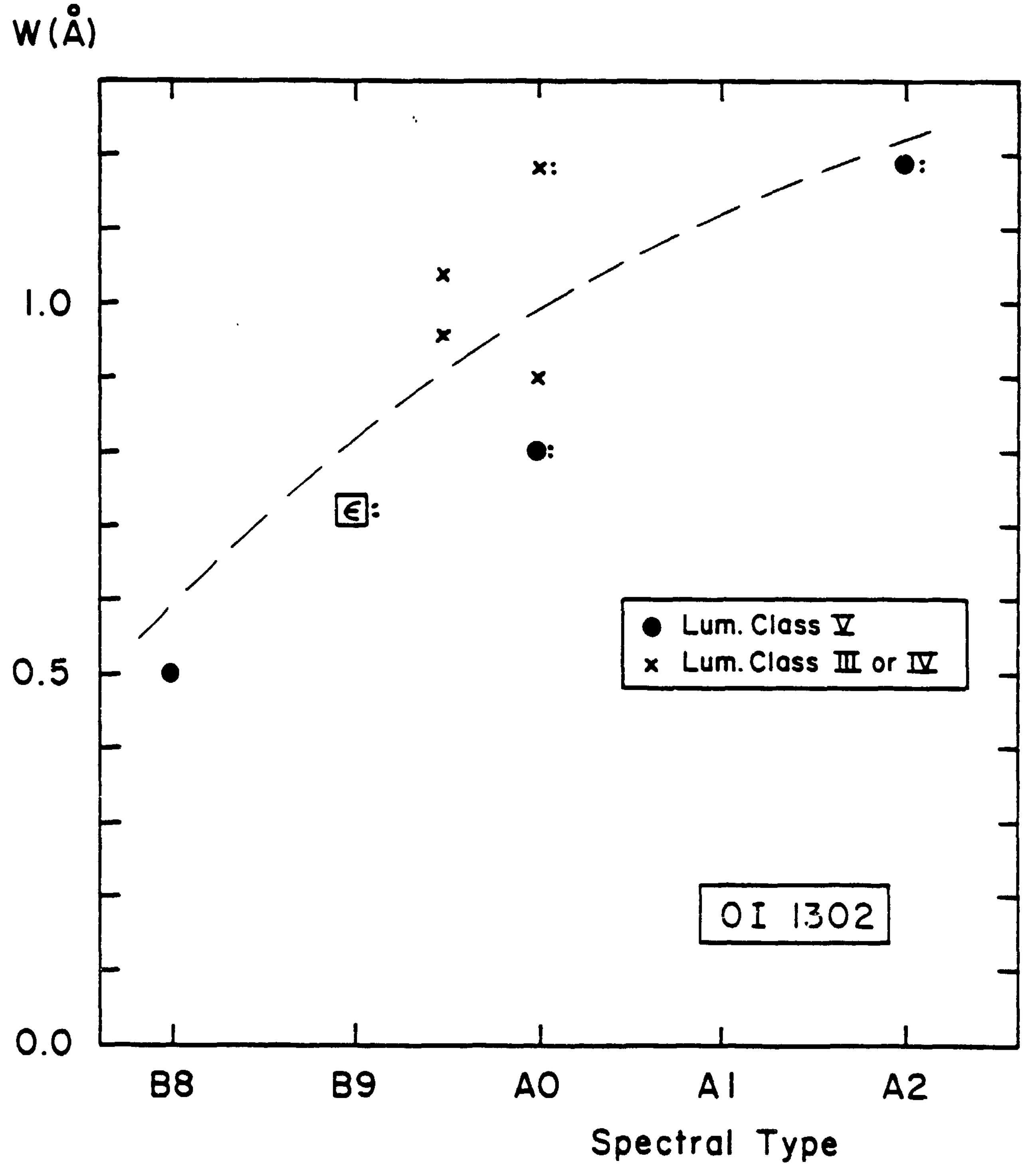

W(Å)
1.0
0.5
0.0
B8
B9
A0
A1
A2
Spectral Type
ϵ:
● Lum. Class V
× Lum. Class III or IV
OI 1302

FIG. 3

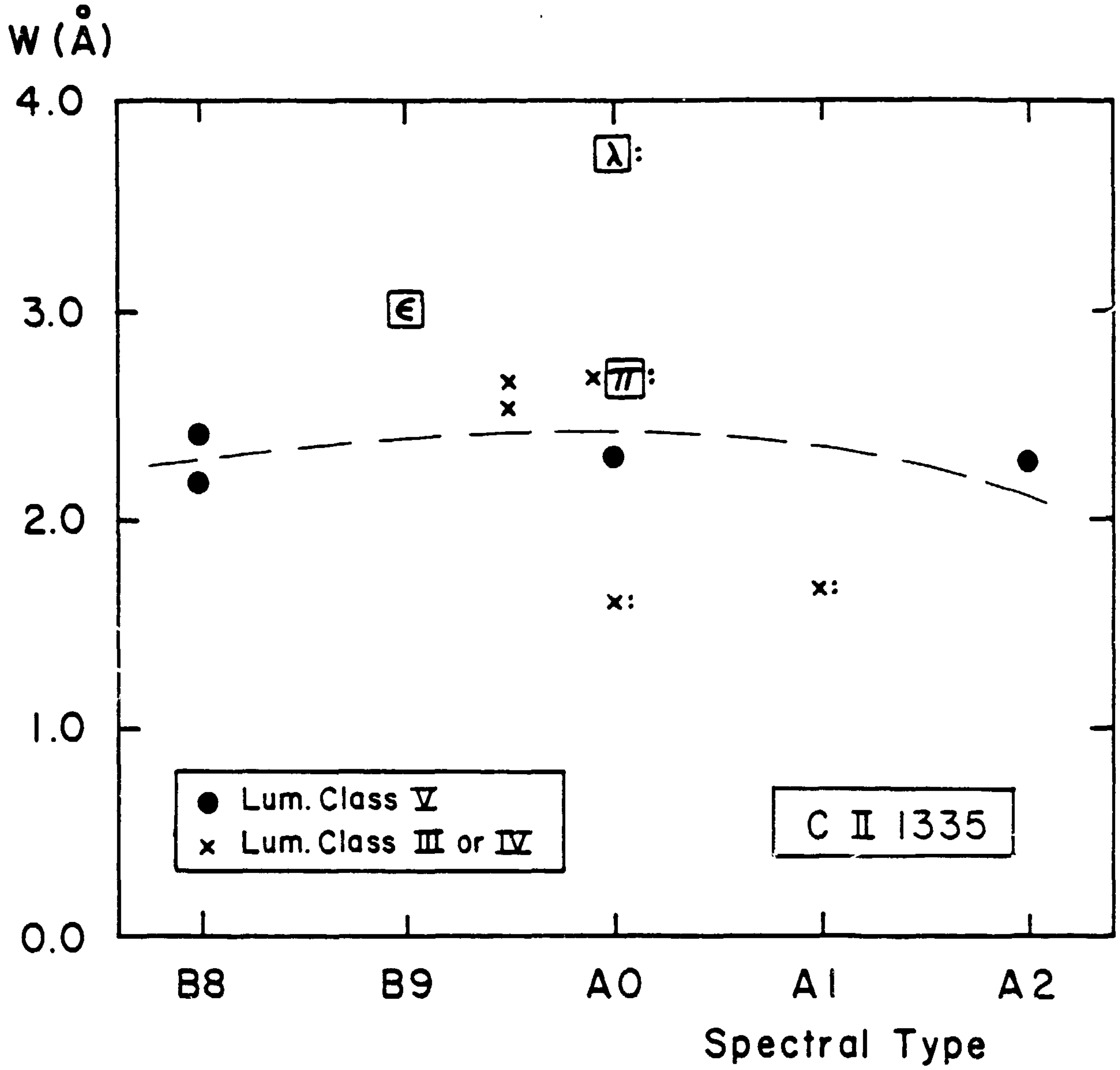

W (Å)
4.0
3.0
2.0
1.0
0.0
λ:
ε
π:
x:
x:
● Lum. Class V
× Lum. Class III or IV
C II 1335
B8
B9
A0
A1
A2
Spectral Type

FIG. 4

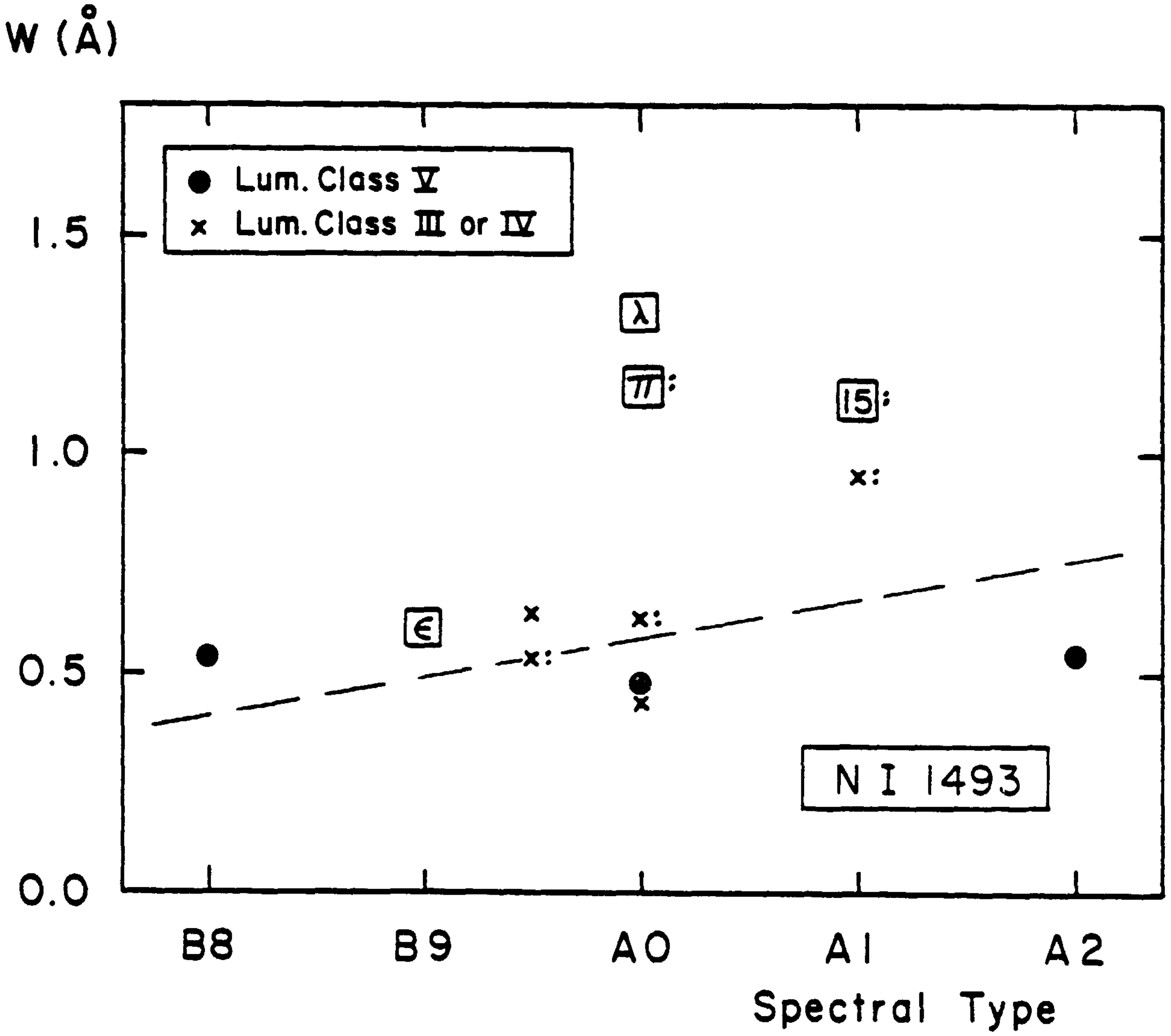

W (Å)
1.5
1.0
0.5
0.0
● Lum. Class V
× Lum. Class III or IV
λ
π :
15 :
×:
ϵ
×
×:
×:
N I 1493
B8
B9
A0
A1
A2
Spectral Type

Fig. 5

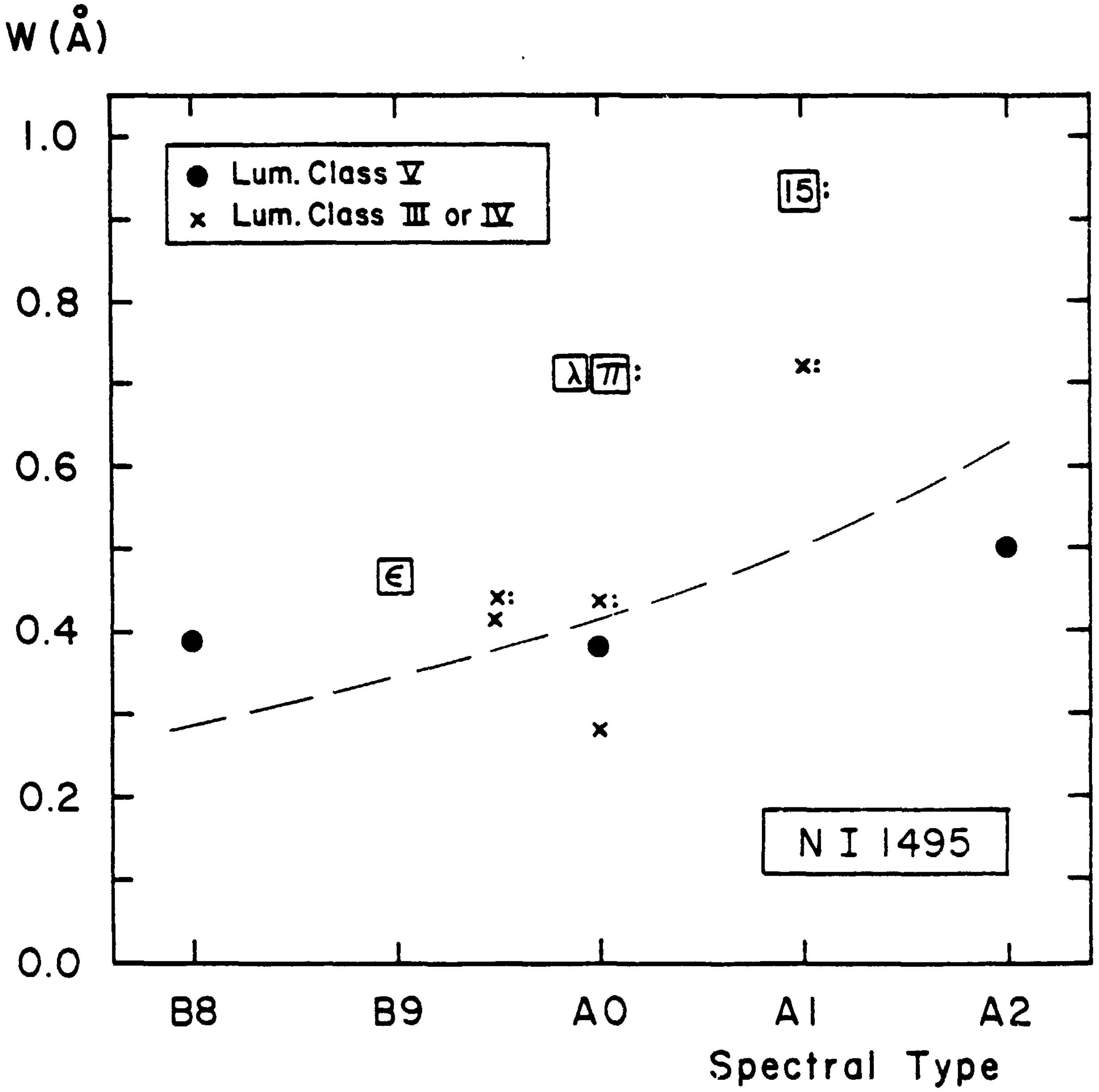

W (Å)
1.0
0.8
0.6
0.4
0.2
0.0
● Lum. Class V
× Lum. Class III or IV
15:
λ π:
ε
×:
×:
×:
N I 1495
B8
B9
A0
A1
A2
Spectral Type

Fig. 6

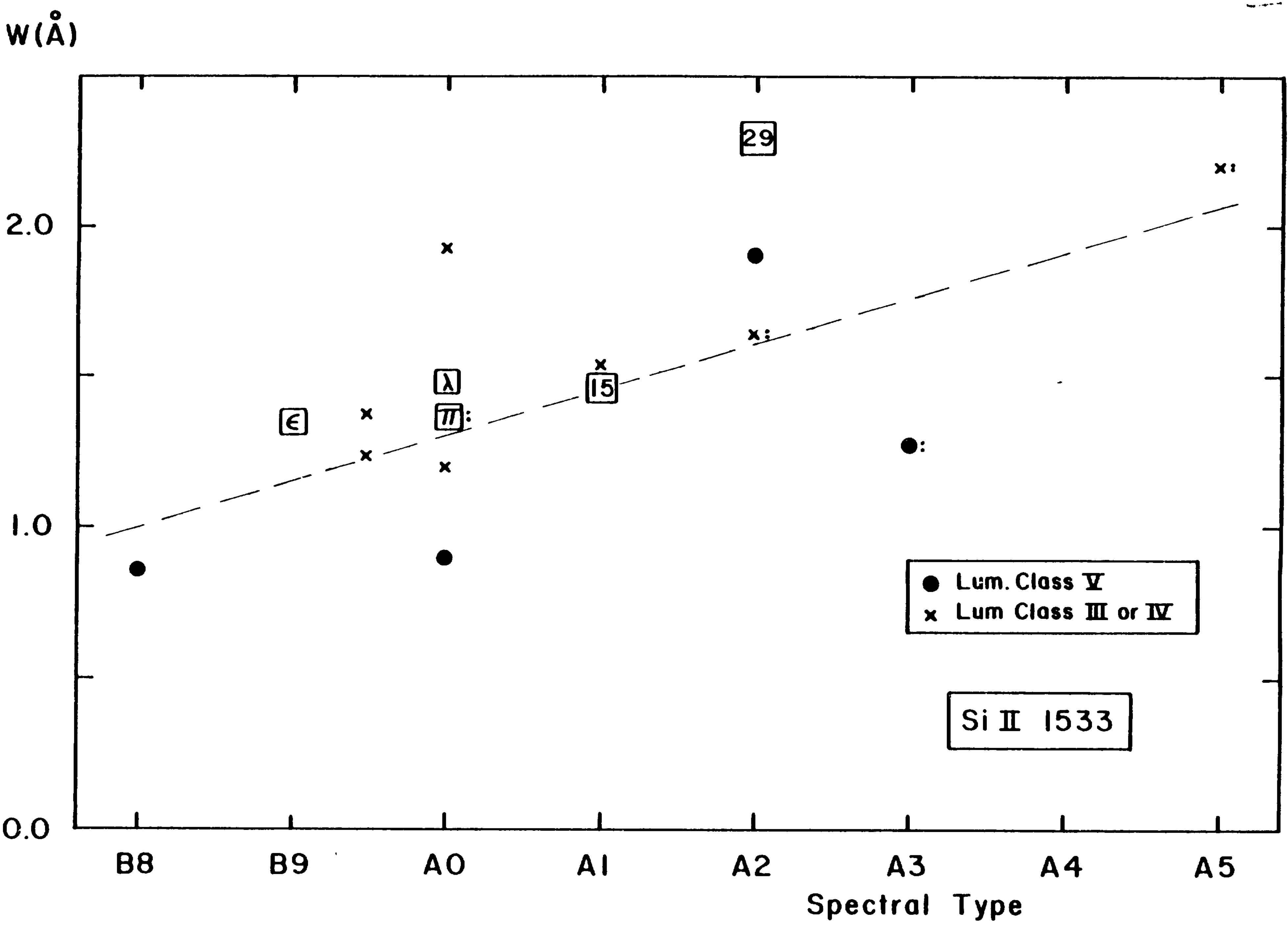

W(Å)
2.0
1.0
0.0
B8
B9
A0
A1
A2
A3
A4
A5
Spectral Type
29
15
λ
π:
ϵ
● Lum. Class V
x Lum Class III or IV
Si II 1533

FIG. 7

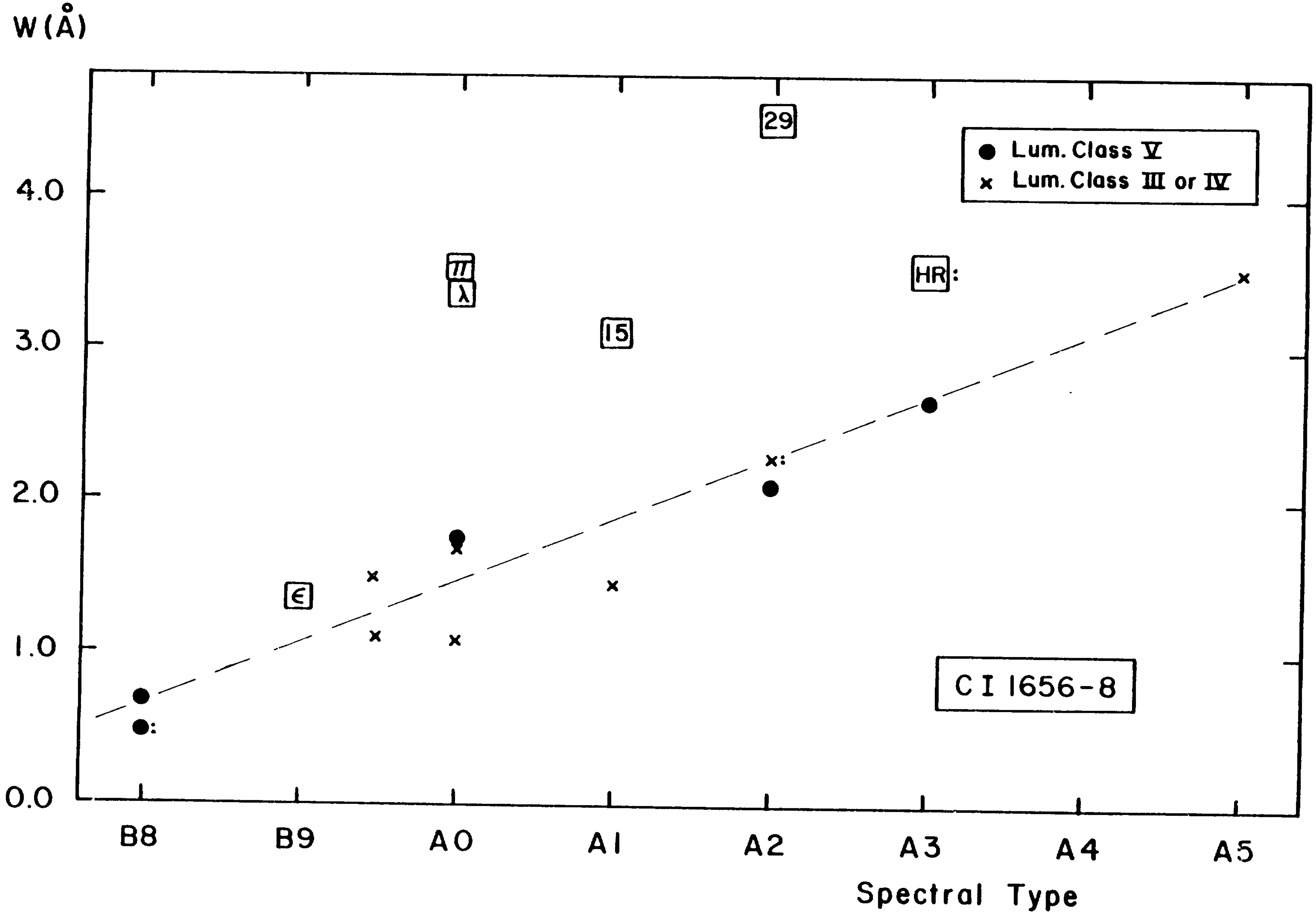
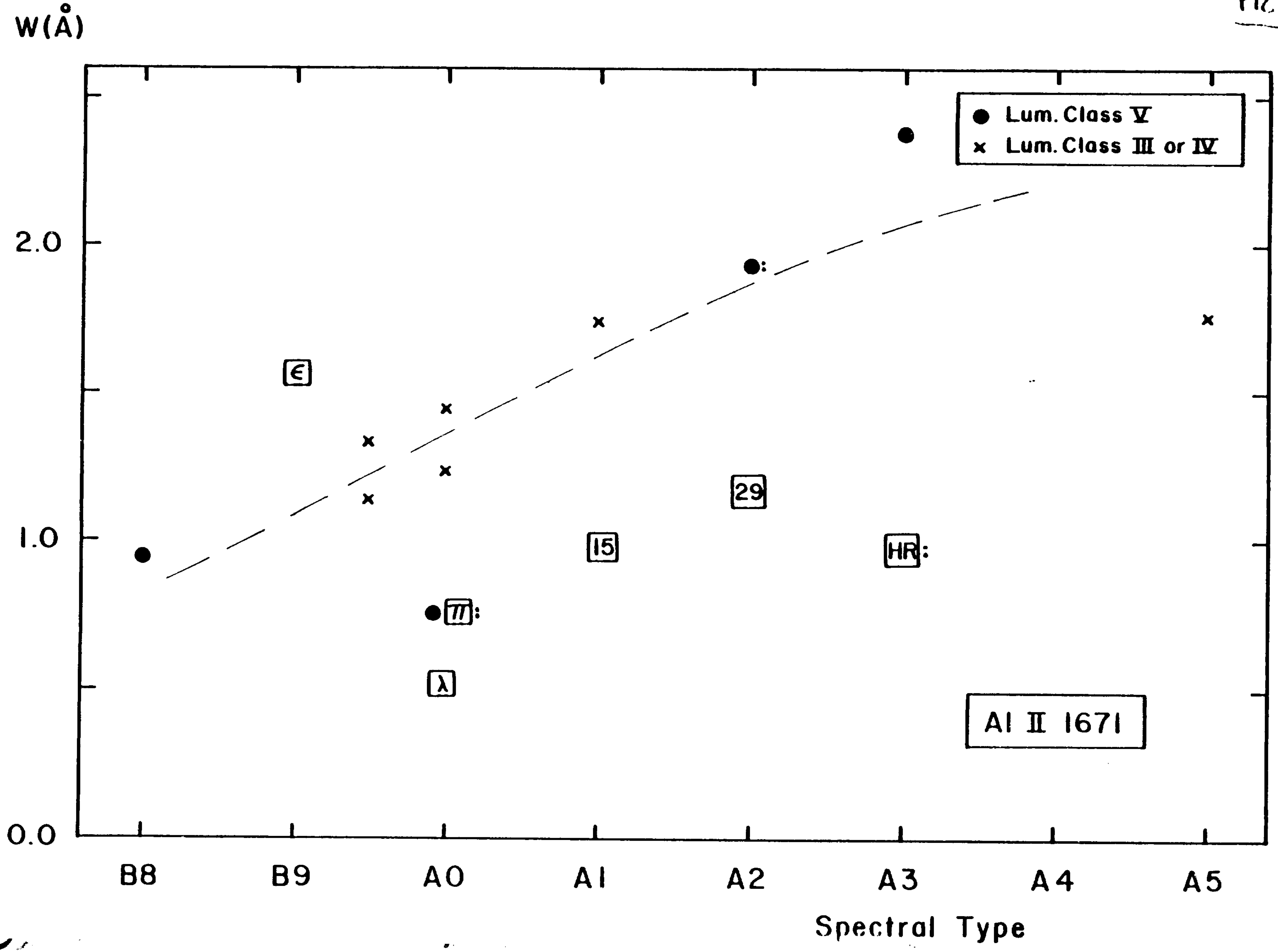

W(Å)
4.0
3.0
2.0
1.0
0.0
B8
B9
A0
A1
A2
A3
A4
A5
Spectral Type
● Lum. Class V
× Lum. Class III or IV
C I 1656-8
29
π
λ
15
HR:
ϵ

Fig. 8


W(Å)
2.0
1.0
0.0
B8
B9
A0
A1
A2
A3
A4
A5
Spectral Type
● Lum. Class V
× Lum. Class III or IV
ε
π:
λ
15
29
HR:
Al II 1671

11([illegible])

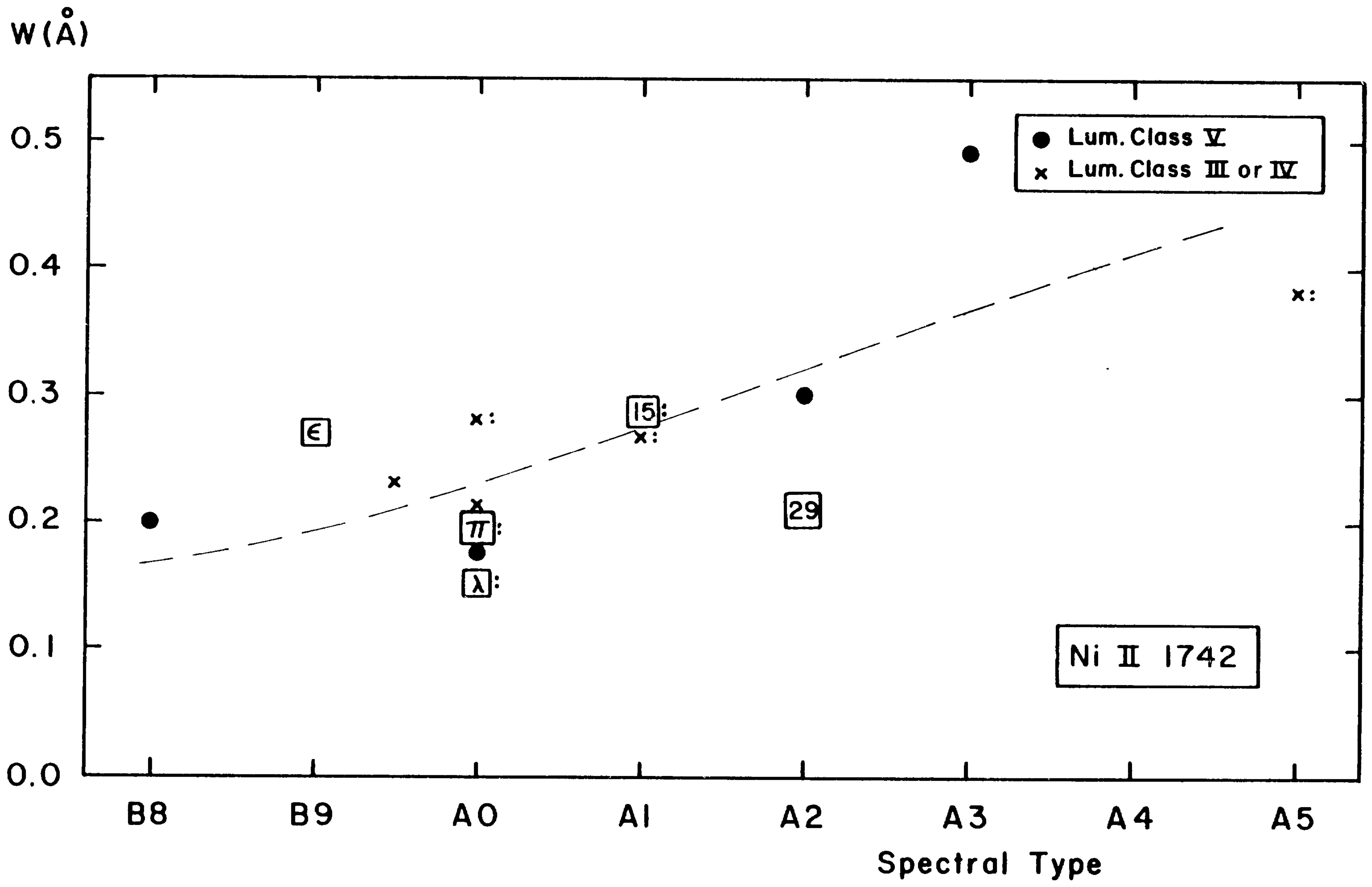

W(Å)
0.5
0.4
0.3
0.2
0.1
0.0
B8
B9
A0
A1
A2
A3
A4
A5
Spectral Type
● Lum. Class V
× Lum. Class III or IV
Ni II 1742
ϵ
π:
λ:
15:
29
×:

FIG. 10

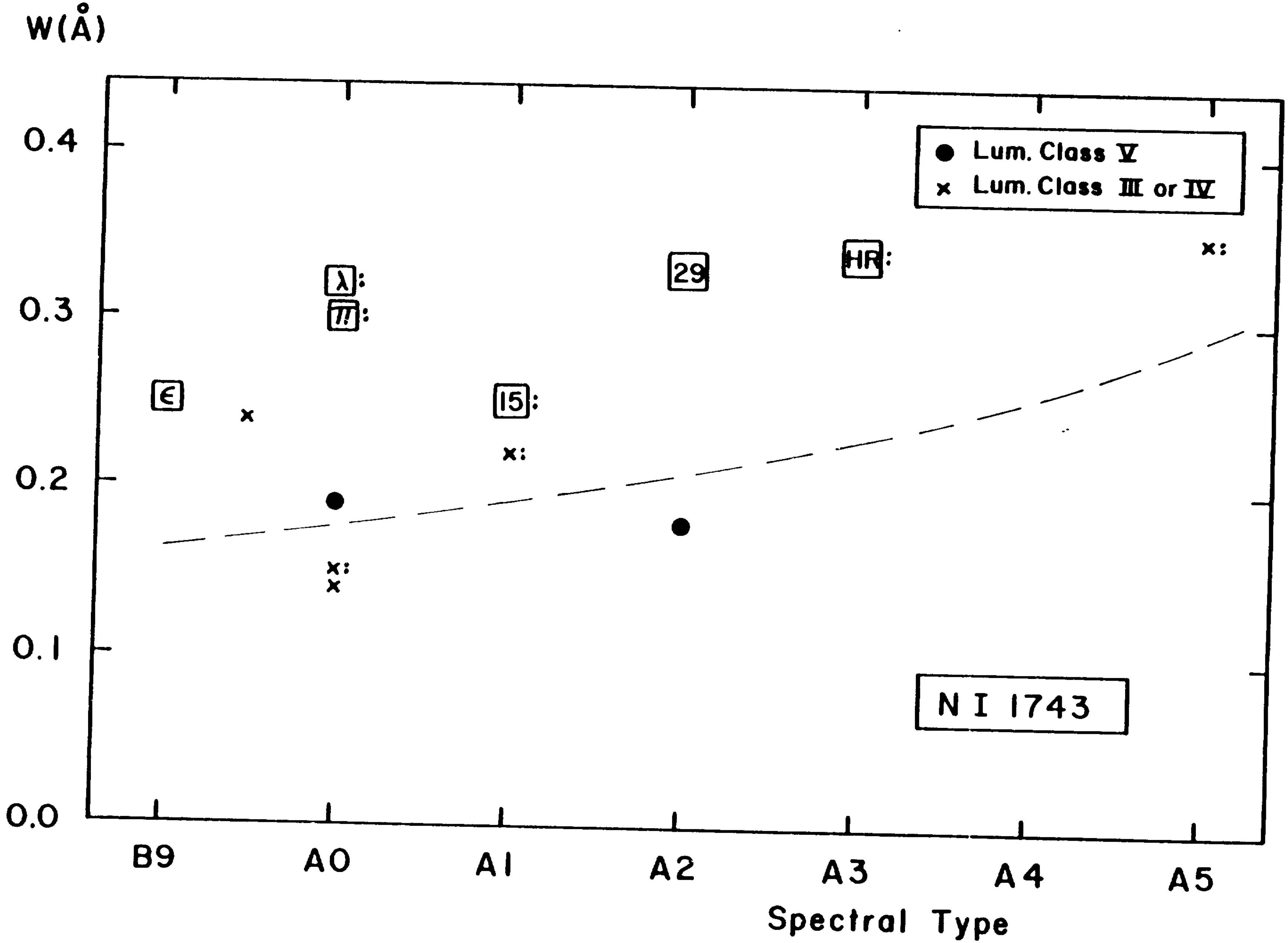

W(Å)
0.4
0.3
0.2
0.1
0.0
B9
A0
A1
A2
A3
A4
A5
Spectral Type
● Lum. Class V
× Lum. Class III or IV
N I 1743
ϵ
λ:
π:
15:
29
HR:
×:

Fig. 11

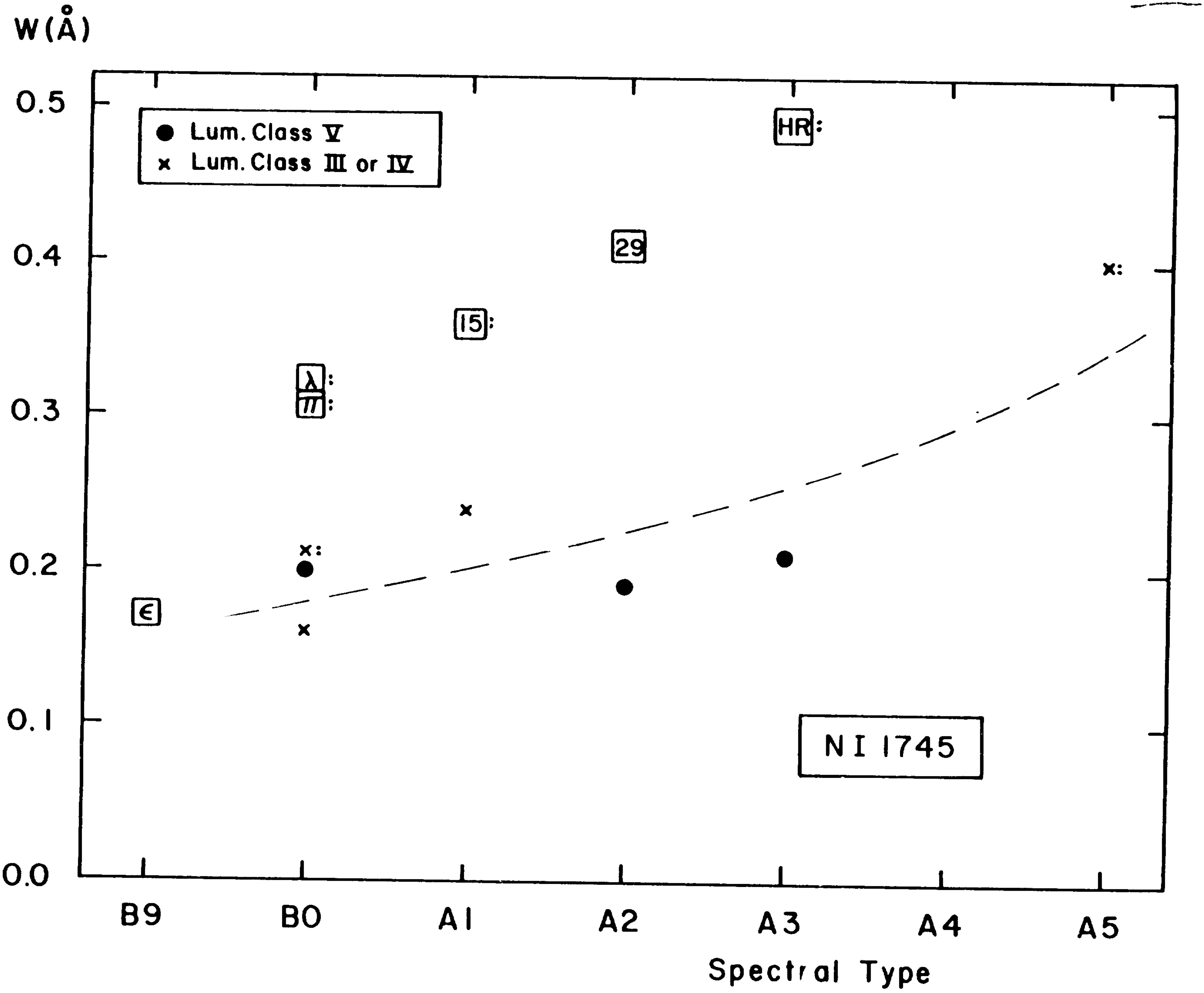

W(Å)
0.5
0.4
0.3
0.2
0.1
0.0
● Lum. Class V
× Lum. Class III or IV
HR:
29
15:
λ:
π:
x:
x:
ε
N I 1745
B9
B0
A1
A2
A3
A4
A5
Spectral Type

Fig. 12

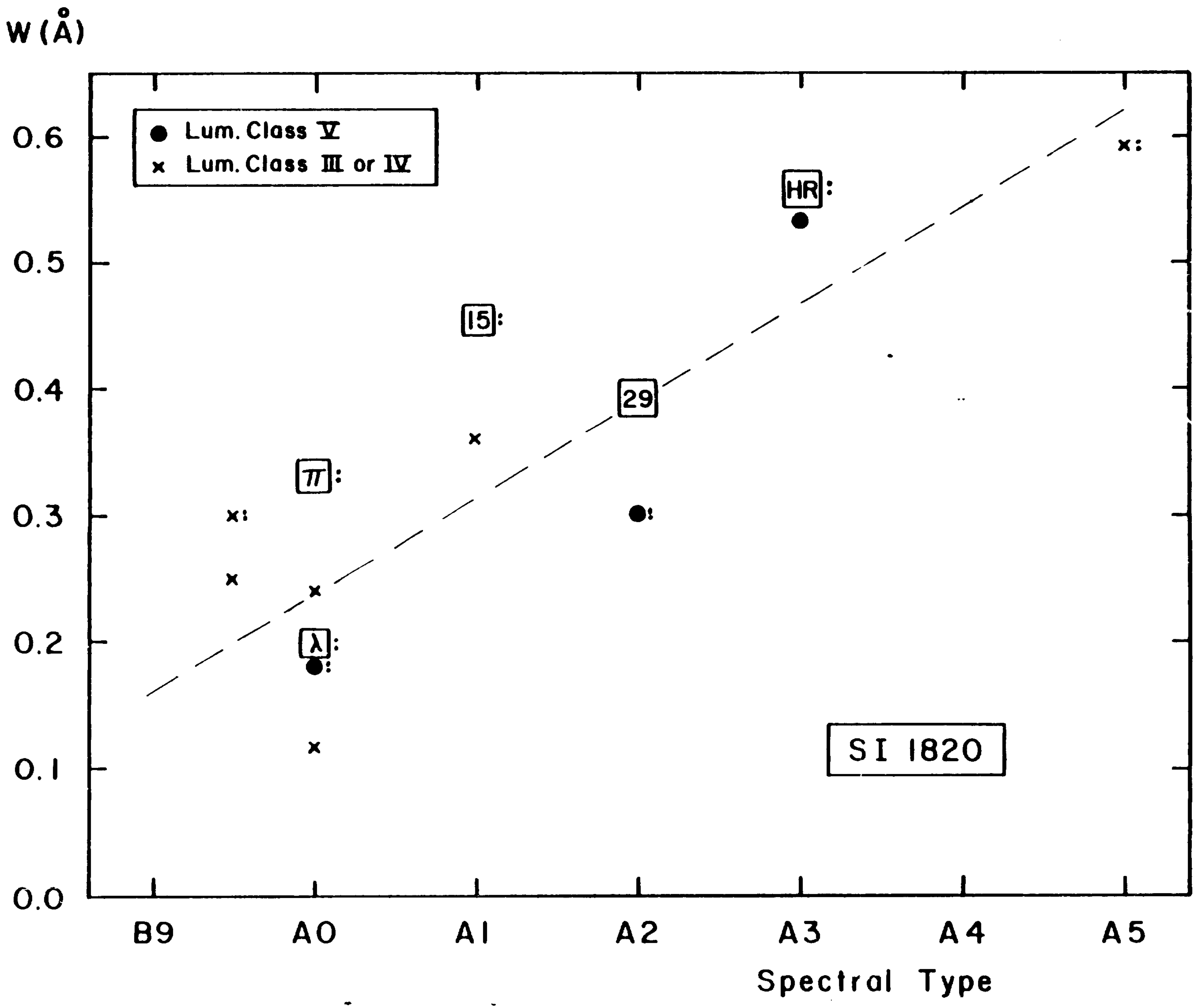

W (Å)
0.6
0.5
0.4
0.3
0.2
0.1
0.0
● Lum. Class V
× Lum. Class III or IV
HR:
15:
29
π:
λ:
SI 1820
B9
A0
A1
A2
A3
A4
A5
Spectral Type

FIG. 13

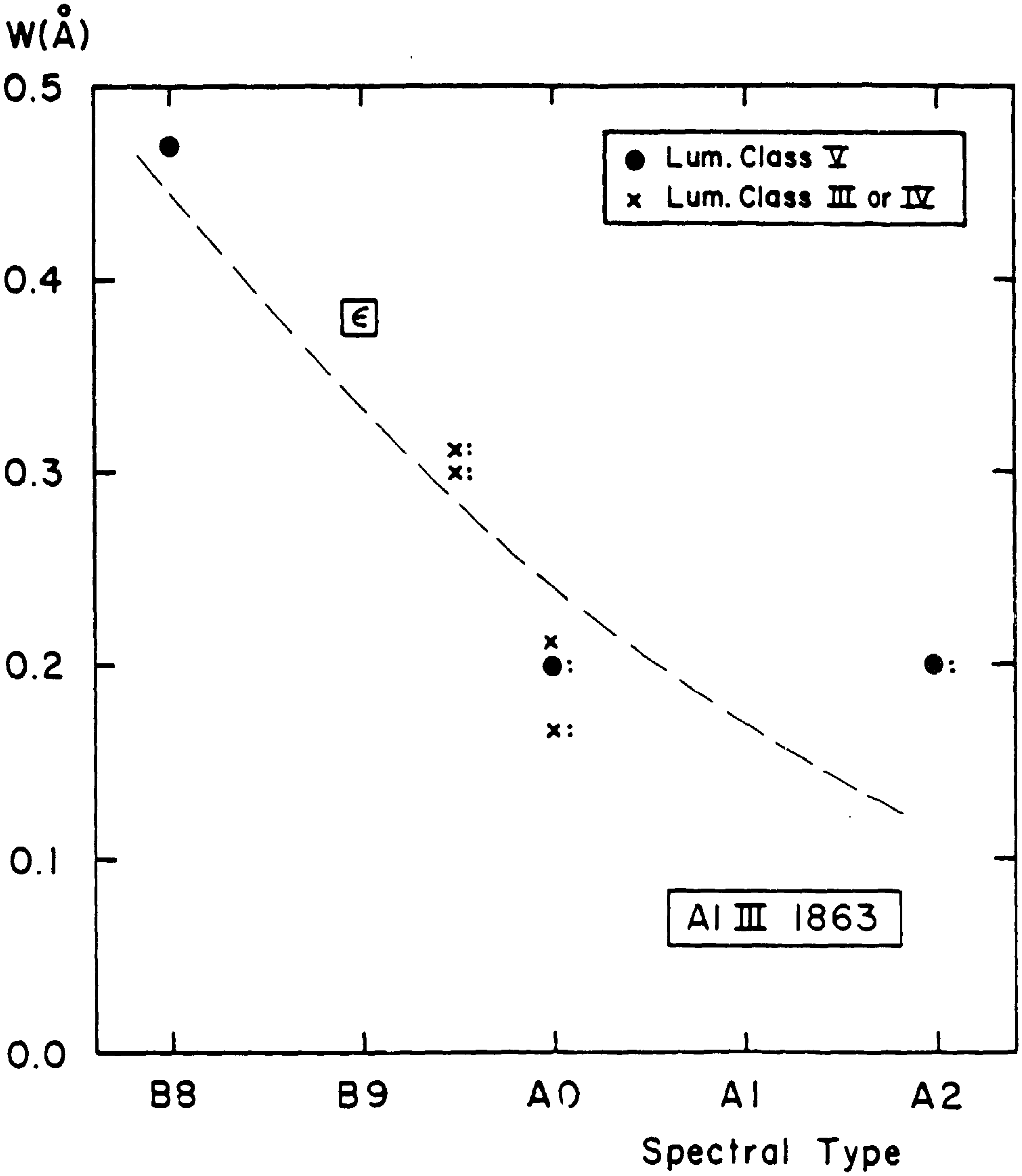

W(Å)
0.5
0.4
0.3
0.2
0.1
0.0
● Lum. Class V
x Lum. Class III or IV
ε
x:
x:
x
●:
x:
●:
Al III 1863
B8
B9
A0
A1
A2
Spectral Type

Fig. 12

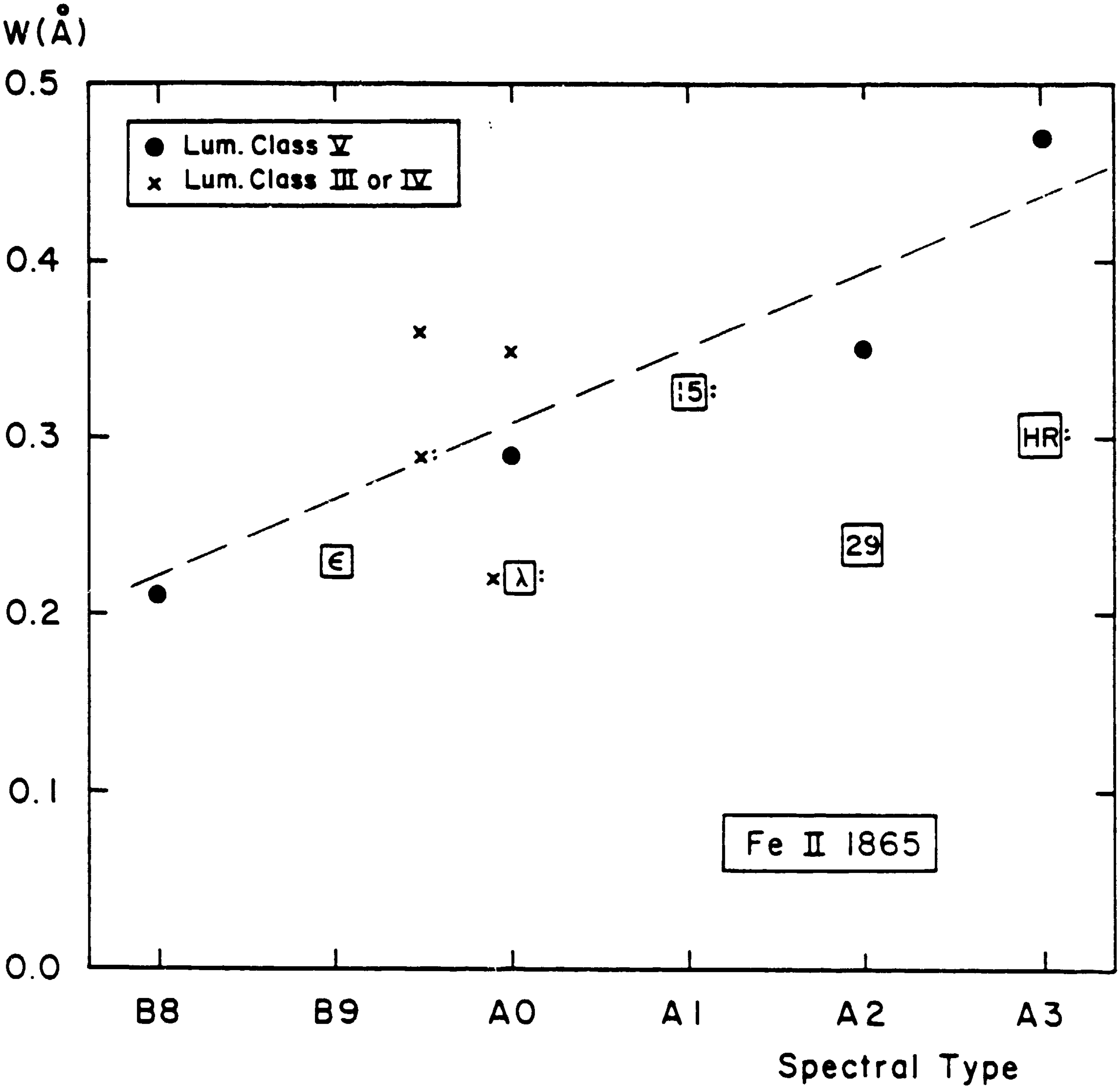
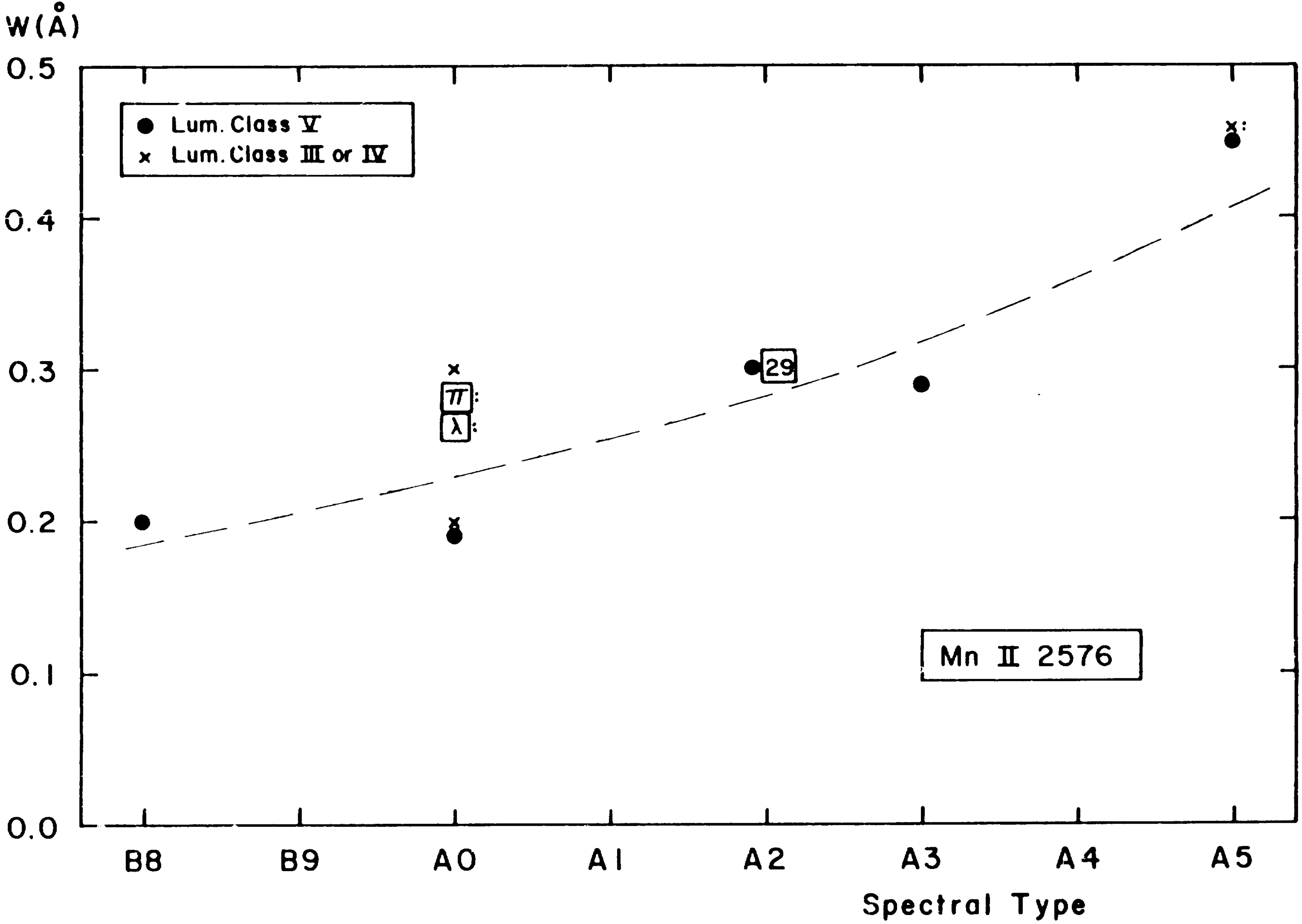

W(Å)
0.5
0.4
0.3
0.2
0.1
0.0
● Lum. Class V
× Lum. Class III or IV
x:
x:
ε
λ:
15:
29
HR:
Fe II 1865
B8
B9
A0
A1
A2
A3
Spectral Type

Fig. 15



W(Å)
0.5
0.4
0.3
0.2
0.1
0.0
● Lum. Class V
× Lum. Class III or IV
π :
λ :
29
x:
Mn II 2576
B8
B9
A0
A1
A2
A3
A4
A5
Spectral Type

Fig. 16

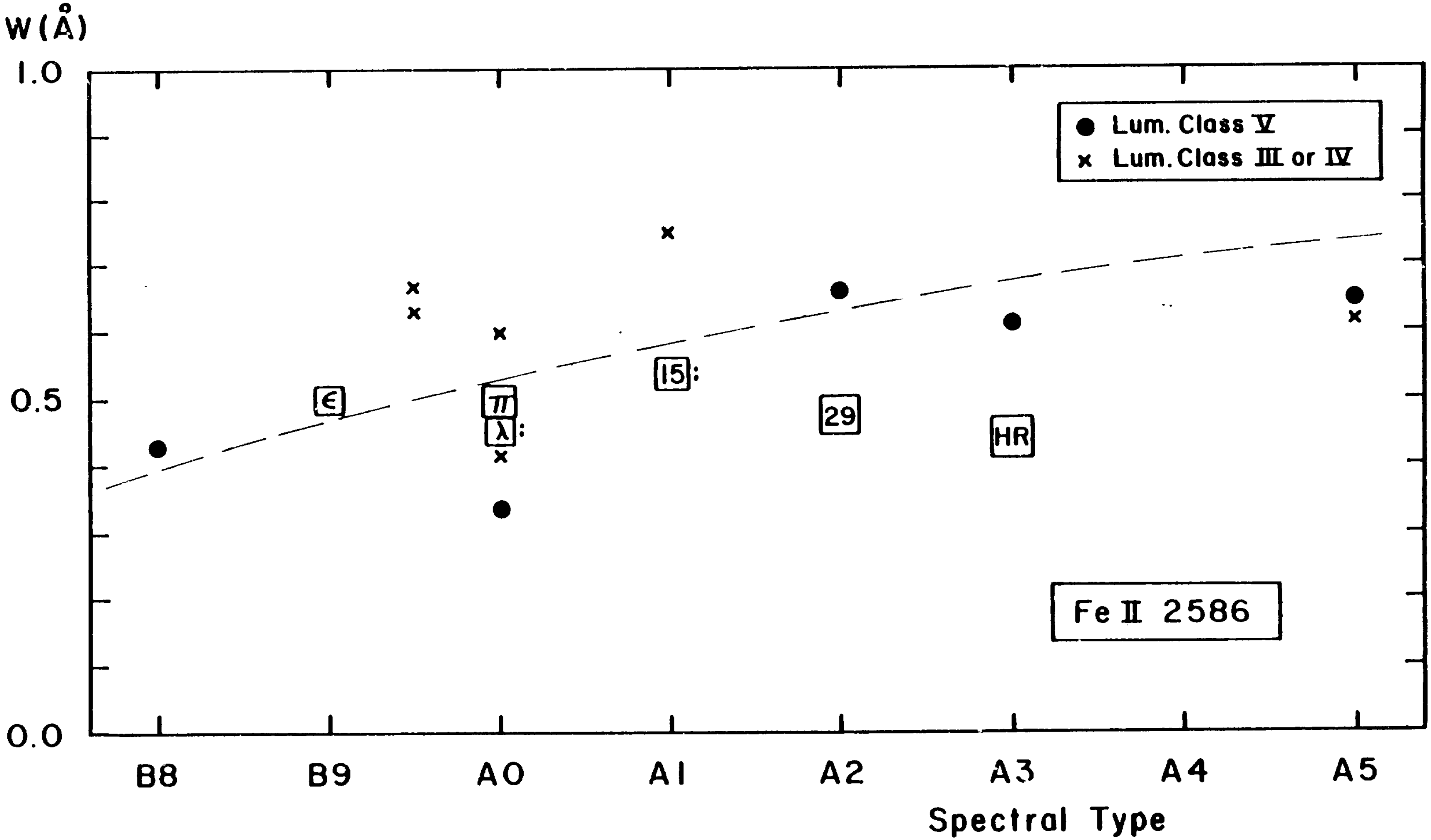

W(Å)
1.0
0.5
0.0
B8
B9
A0
A1
A2
A3
A4
A5
Spectral Type
● Lum. Class V
× Lum. Class III or IV
ε
π
λ:
15:
29
HR
Fe II 2586

Fig. 11

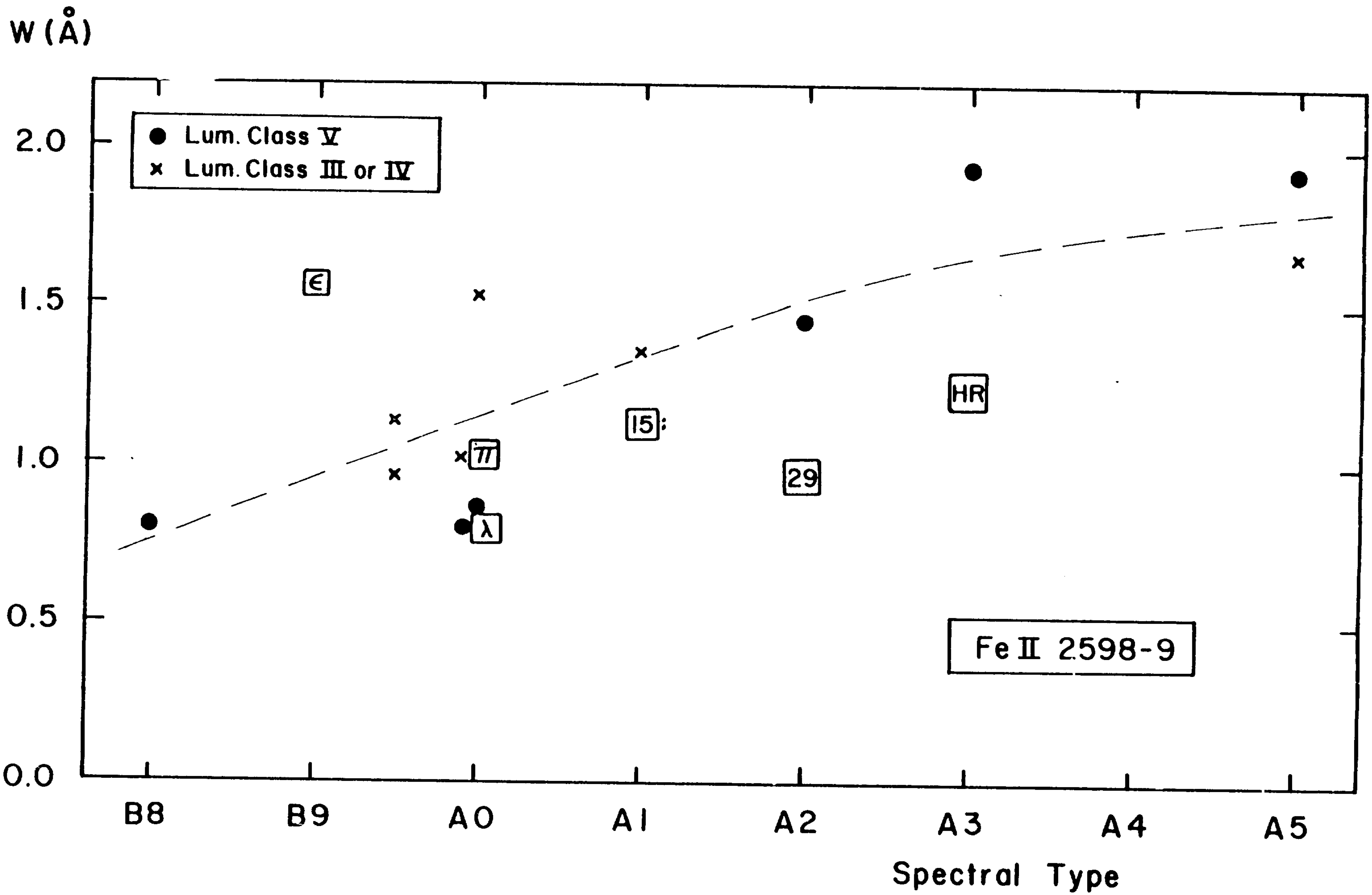

W (Å)
2.0
1.5
1.0
0.5
0.0
● Lum. Class V
× Lum. Class III or IV
ε
π
λ
15:
29
HR
Fe II 2598-9
B8
B9
A0
A1
A2
A3
A4
A5
Spectral Type

Fig. 11

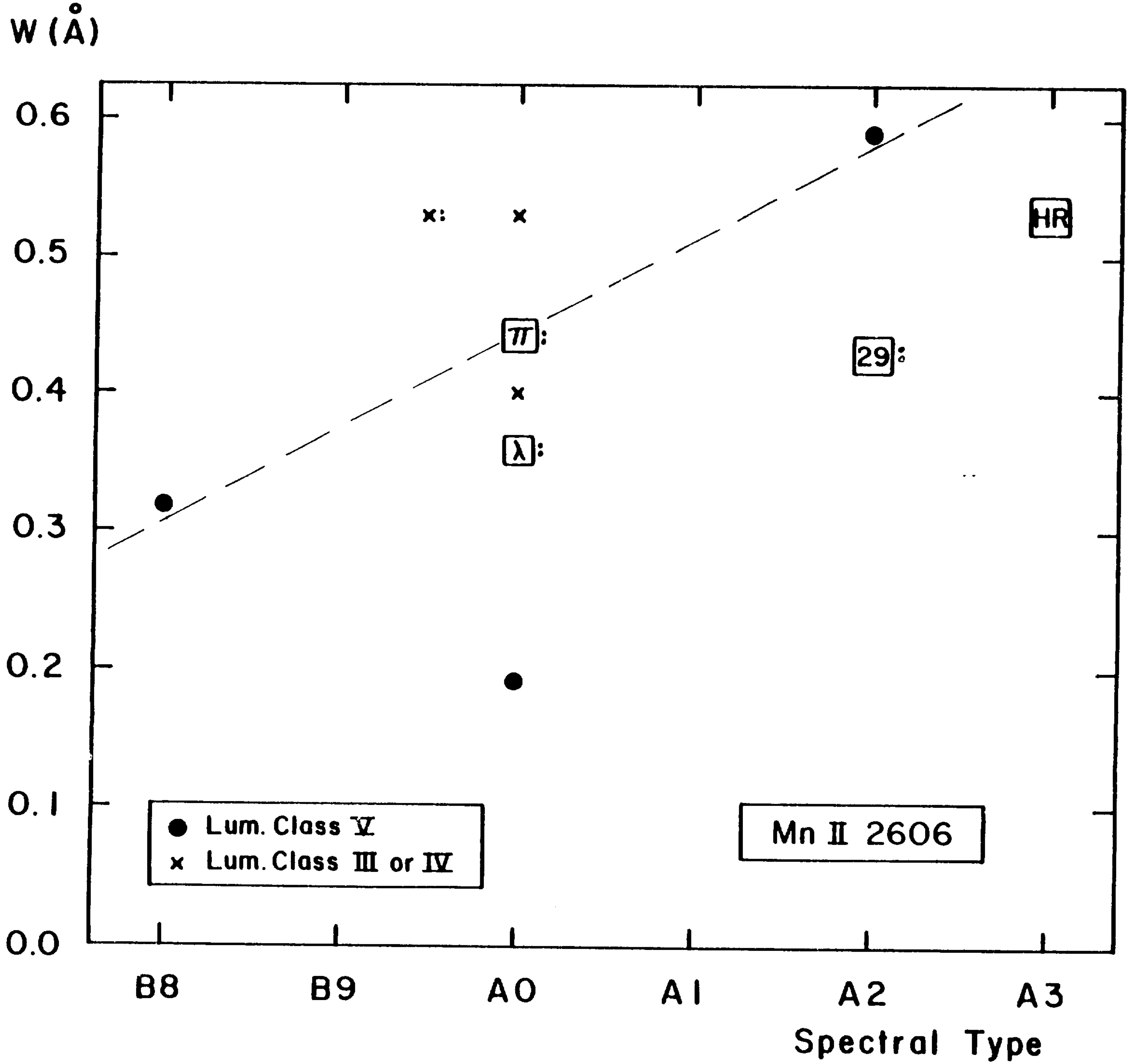
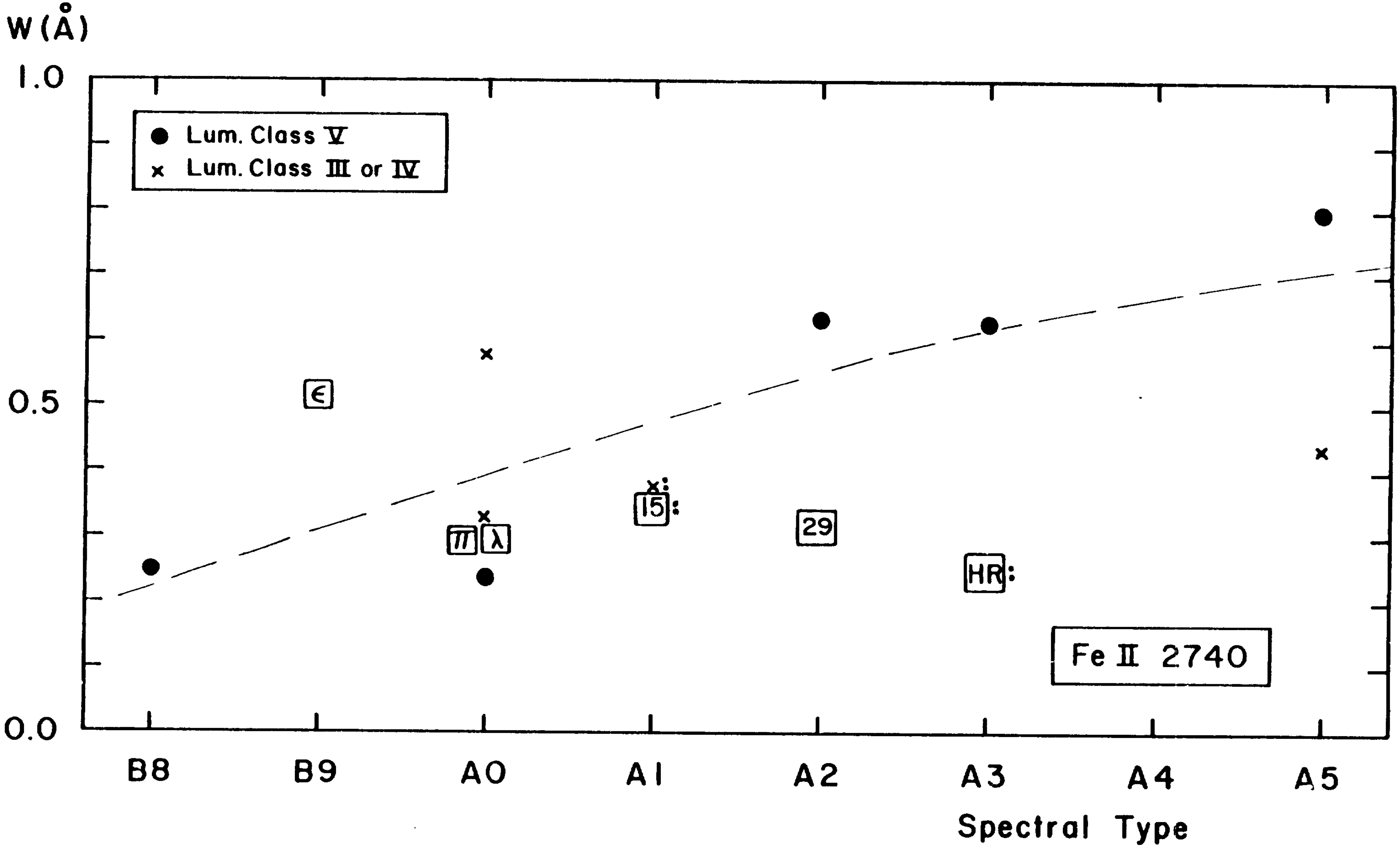

W (Å)
0.6
0.5
0.4
0.3
0.2
0.1
0.0
B8
B9
A0
A1
A2
A3
Spectral Type
x:
x
π:
x
λ:
29:
HR
● Lum. Class V
x Lum. Class III or IV
Mn II 2606

FIG. [illegible]


W (Å)
1.0
0.5
0.0
● Lum. Class V
× Lum. Class III or IV
ϵ
π
λ
15
29
HR
Fe II 2740
B8
B9
A0
A1
A2
A3
A4
A5
Spectral Type

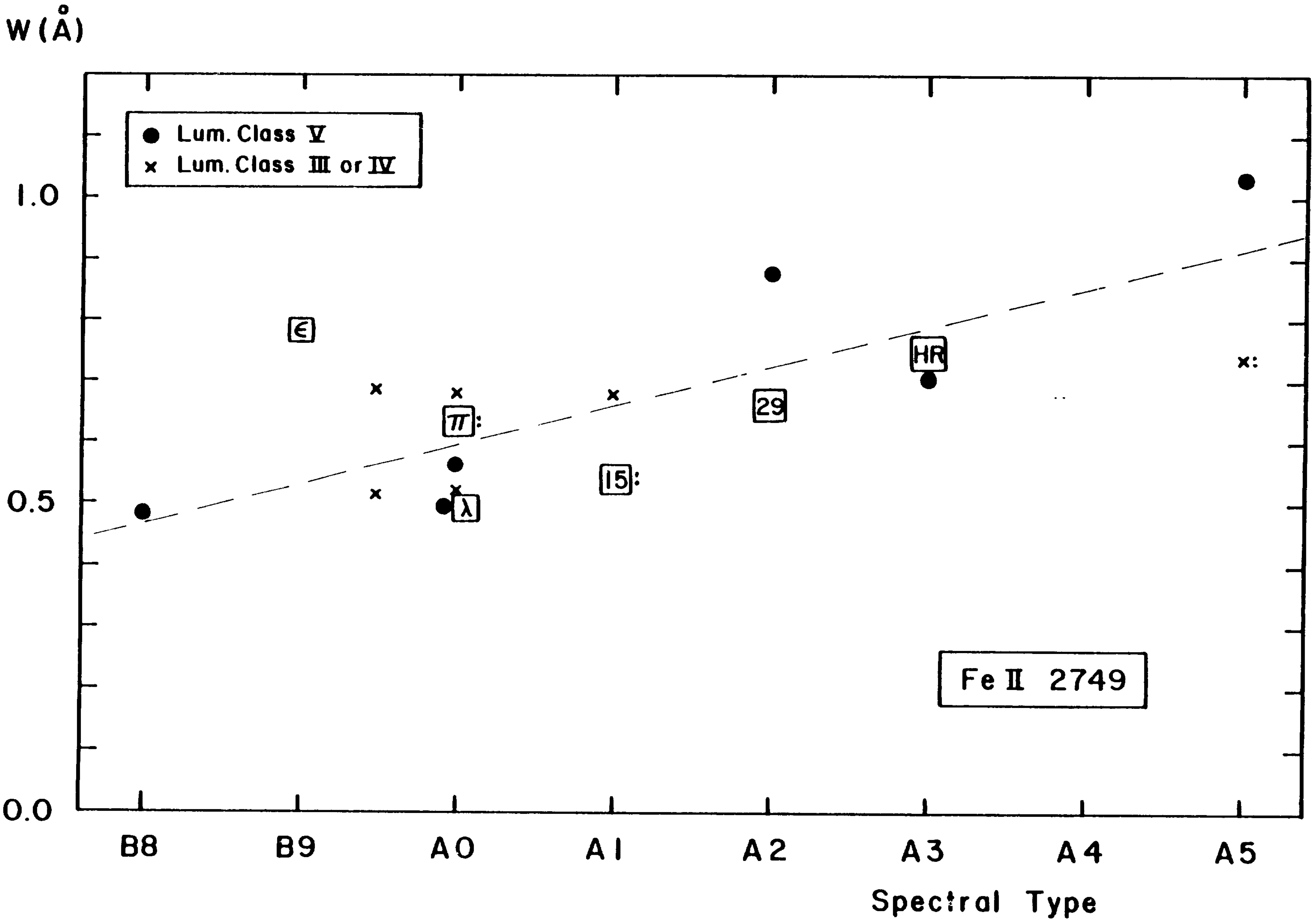

W(Å)
1.0
0.5
0.0
Lum. Class V
Lum. Class III or IV
ε
π:
λ
15:
29
HR
x:
Fe II 2749
B8
B9
A0
A1
A2
A3
A4
A5
Spectral Type

FIG. 21

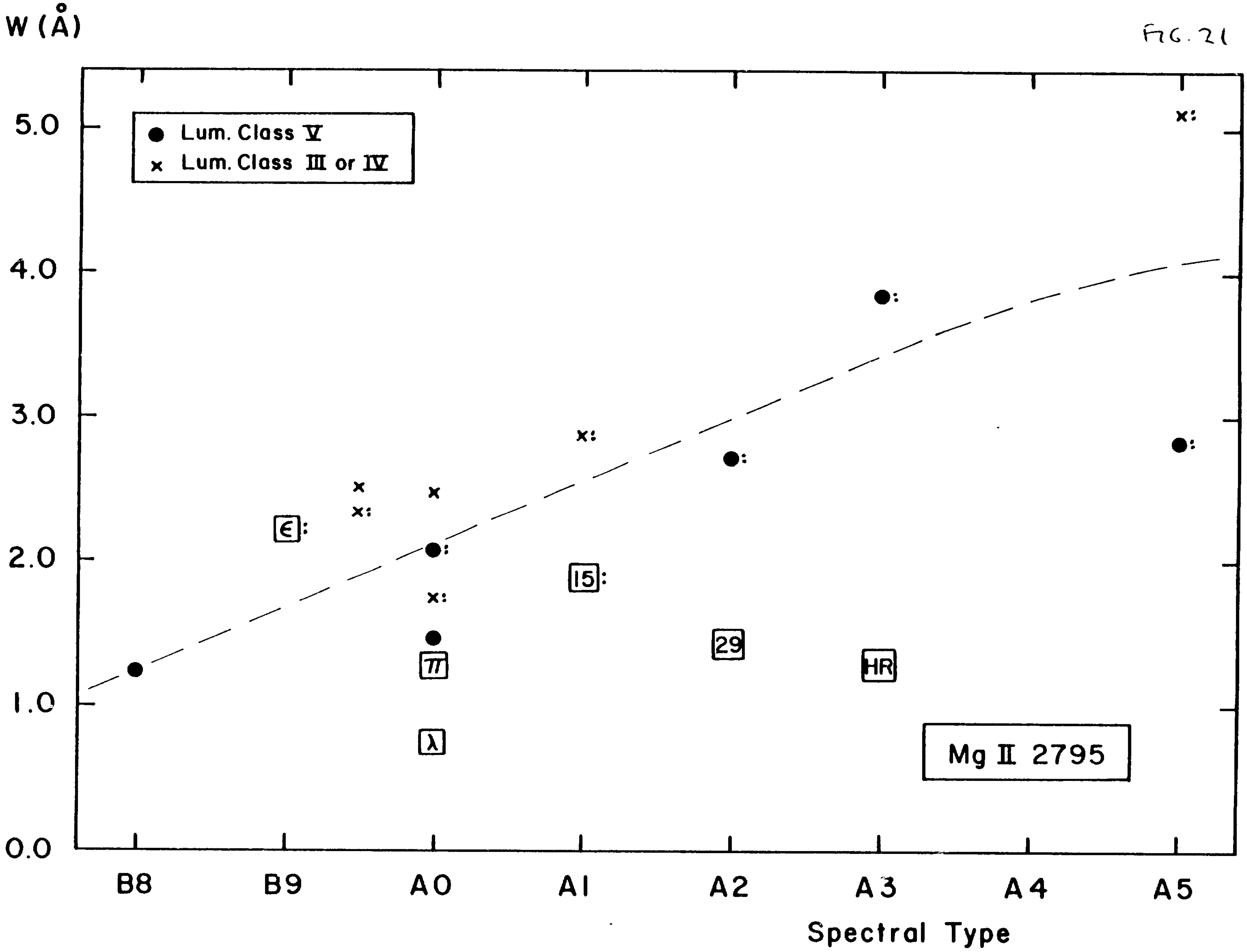

W (Å)
5.0
4.0
3.0
2.0
1.0
0.0
● Lum. Class V
x Lum. Class III or IV
ε:
π
λ
15:
29
HR
Mg II 2795
B8
B9
A0
A1
A2
A3
A4
A5
Spectral Type

FIG. 12

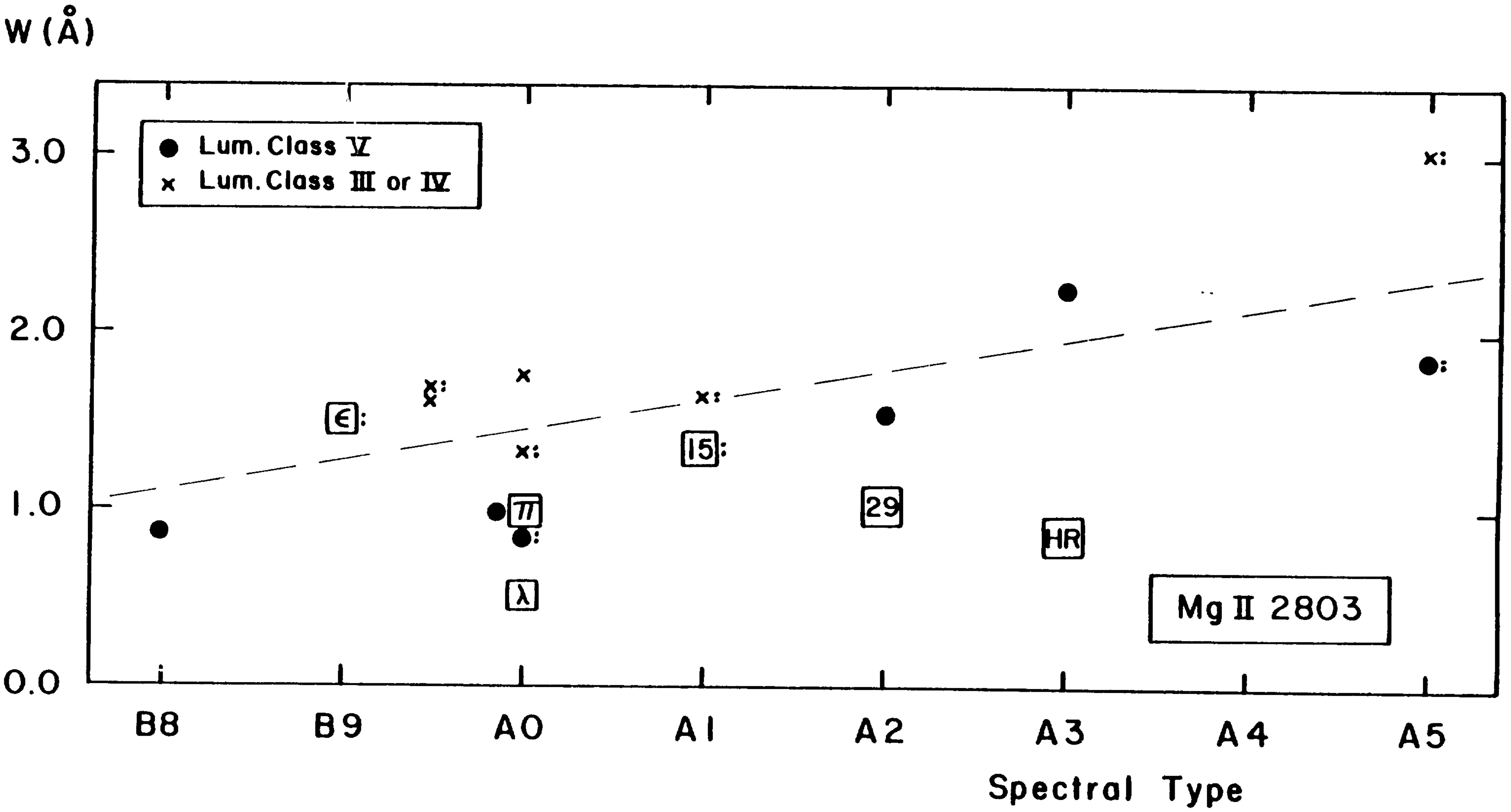

W (Å)
3.0
2.0
1.0
0.0
● Lum. Class V
× Lum. Class III or IV
ϵ:
x:
x
x:
π
•:
λ
x:
15:
29
HR
x:
•:
Mg II 2803
B8
B9
A0
A1
A2
A3
A4
A5
Spectral Type

FIG. 13

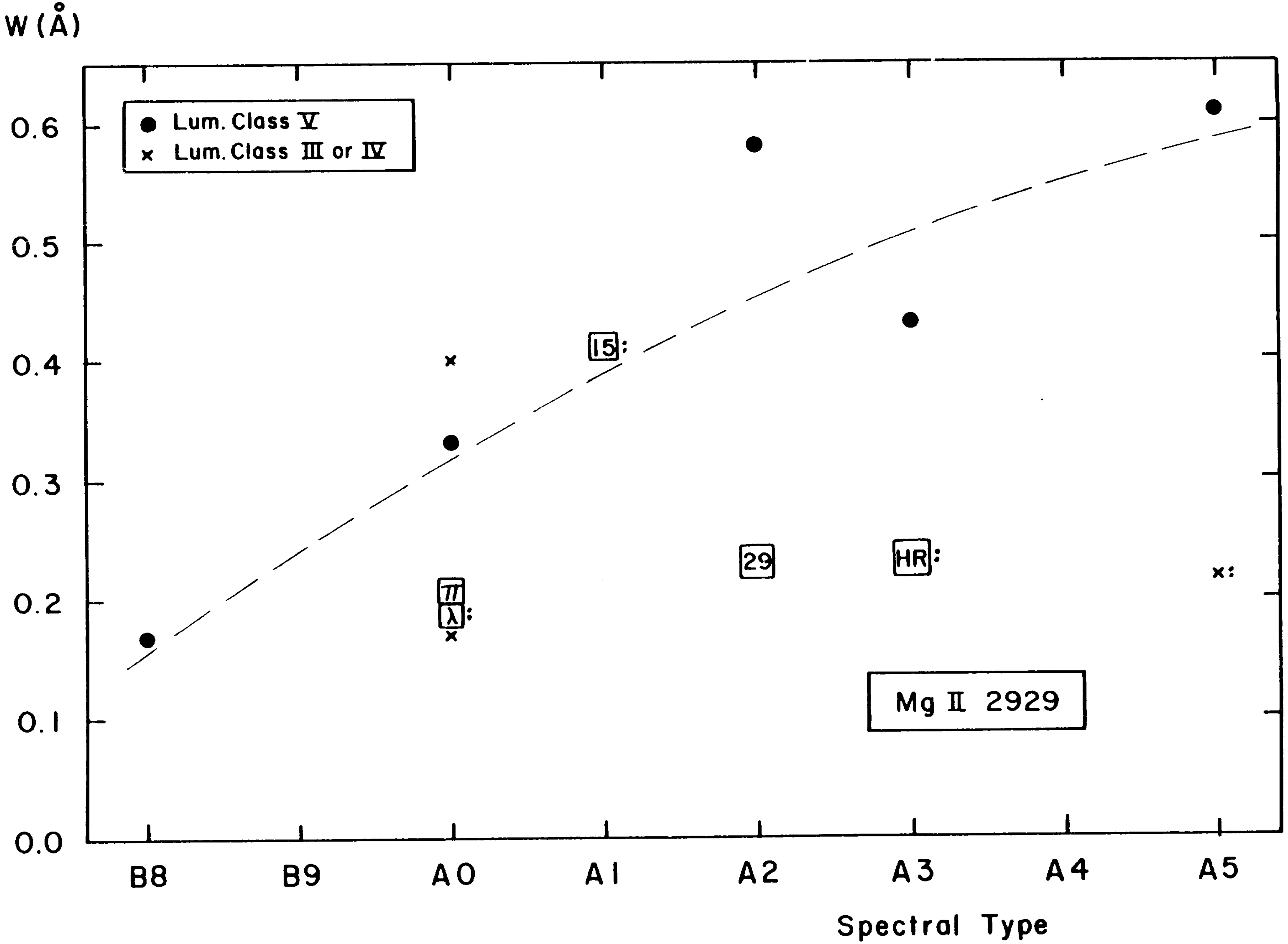
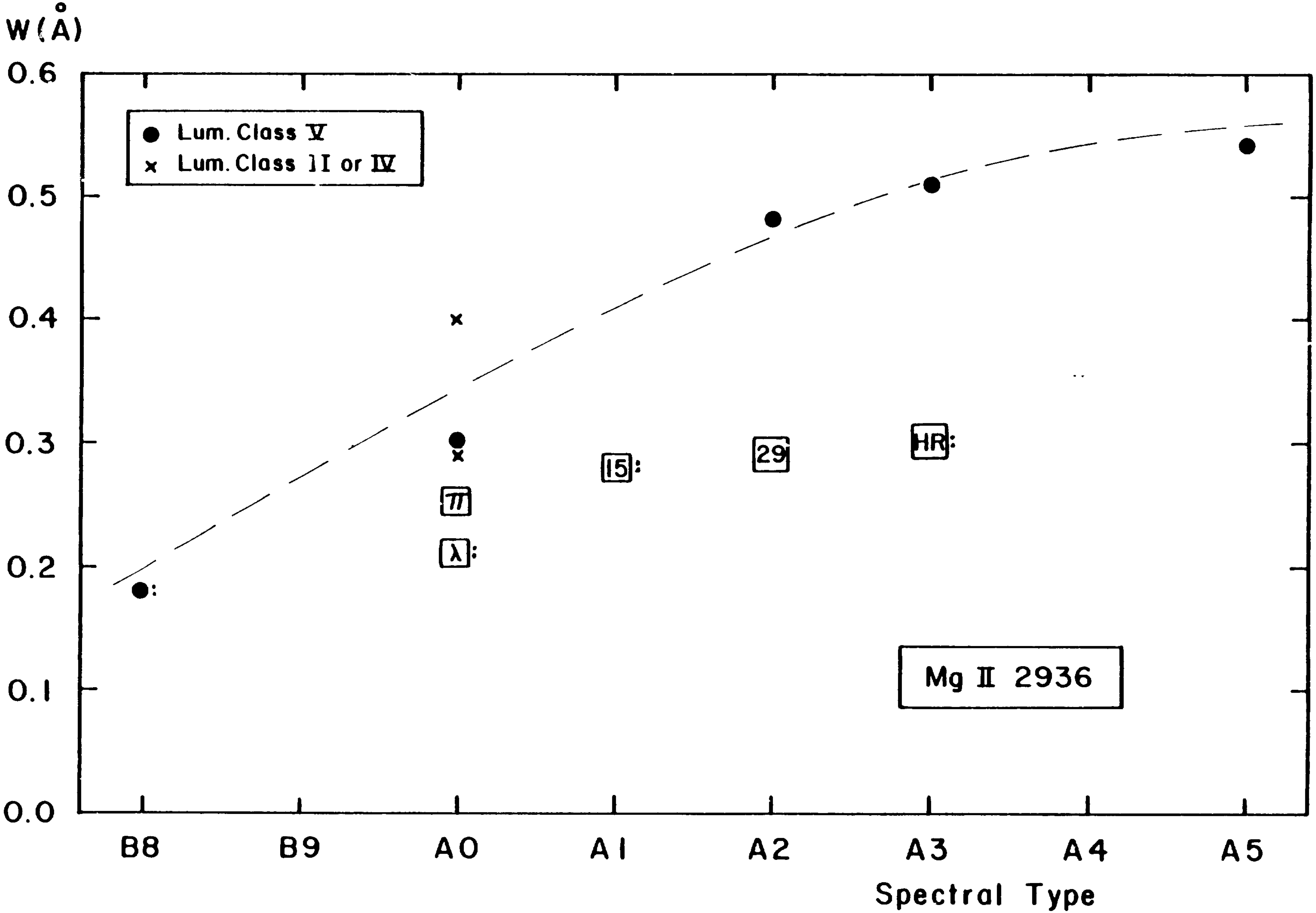

W (Å)
0.0
0.1
0.2
0.3
0.4
0.5
0.6
● Lum. Class V
× Lum. Class III or IV
15:
29
HR:
π
λ:
×:
Mg II 2929
B8
B9
A0
A1
A2
A3
A4
A5
Spectral Type

FIG. 21


W(Å)
0.6
0.5
0.4
0.3
0.2
0.1
0.0
● Lum. Class V
× Lum. Class II or IV
15:
29
HR:
π
λ:
Mg II 2936
B8
B9
A0
A1
A2
A3
A4
A5
Spectral Type

END

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