

FAR-ULTRAVIOLET EMISSION LINES IN U CEPHEI: EVIDENCE FOR A HOT, TURBULENT CIRCUMSTELLAR ENVELOPE

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ABSTRACT

U Cephei is a classical semidetached eclipsing binary of the Algol type, whose unusually high level of activity in the optical spectrum has long been recognized. This paper reports detection of numerous emission lines in the far-ultraviolet spectrum. The emission lines are quite similar to those observed in the *W Serpentis* stars, only weaker and observable only during the primary eclipse. For the first time, the line-emitting region has been studied both in emission and in absorption. It surrounds the hotter component and extends to a distance of several of its radii, although the main part of the emission is coming from the immediate vicinity of the star. A model is proposed in which the emission originates in a circumstellar region at an electron temperature of about 10,000 K. This region is highly turbulent ($v \approx 100 \text{ km s}^{-1}$). The absence of intercombination lines suggests an electron density $n_e \geq 5 \times 10^{11} \text{ cm}^{-3}$, but if this value is correct, the emitting region is probably clumpy and may be formed by several shocked regions. The energy needed for ionization and excitation is derived from collisions and is most likely ultimately extracted from the accretion energy. The circumstellar material, coming from the Roche-filling secondary star, appears to have been at least partially processed by the CNO cycle. The effective temperature of the hotter star is found to be 11,250 K, the distance to the system 208 pc, and the color excess $E(B - V) = 0.03 \text{ mag}$.

Subject headings: stars: circumstellar shells — stars: eclipsing binaries — stars: emission-line — stars: individual — ultraviolet: spectra

1. IMPORTANCE OF THE SYSTEM

Observations with the *International Ultraviolet Explorer* (*IUE*) satellite have led to the discovery that some eclipsing binary stars exhibit very strong emission lines of N v, C iv, Si iv, Fe iii, Al iii, etc., in the far-ultraviolet region (Plavec and Koch 1978). I called this group of binaries the *W Serpentis* stars (Plavec 1980). The ionization energy for these emissions cannot be supplied by the component stars, and evidence is accumulating that they do not originate in the chromospheres of the late-type components. Most likely the required energy ultimately comes from the accretion process, since all the systems of this group show evidence of a large-scale mass transfer. However, the identification of the physical processes involved runs into difficulties because of the nature of these systems. It is evident that the accreting components are surrounded by circumstellar matter which seems to have a complicated structure, so that we can tentatively talk about disks, shells, and stellar wind. This structure complicates the light curves and the radial velocity curves to such a degree that for none of the seven systems of *W Serpentis* type do we have a satisfactory picture about the masses and dimensions. Plavec, Weiland, and Koch (1982) recently analyzed SX Cassiopeiae as the most tractable system, yet even in this object the uncertainty in the radial velocity curves

leaves important questions unanswered. For example, it is not clear if the secondary star actually fills its critical Roche lobe, so that the mode of its mass loss remains uncertain.

It was rather imperative to look for better defined systems with *W Ser* characteristics. The objects recognized so far as members of the *W Serpentis* group were originally selected for *IUE* observations because of the presence of strong emission lines in the optical spectra (mainly the Balmer lines). A fairly large number of the semidetached systems of the Algol type display emission lines (Plavec and Polidan 1976), and these then are candidates for membership in the *W Serpentis* group. U Cephei, because of its unusually high circumstellar activity compared with other "classical" Algol-type eclipsing binaries, was an obvious target. Out-of-eclipse observations, both by myself and by other groups (Kondo, McCluskey, and Stencel 1979; Kondo, McCluskey, and Harvel 1981) failed to show any emission lines in the ultraviolet, except for the Mg ii resonance doublet. However, observations during the primary eclipse, made by me in 1981 August and December did show a rich emission-line spectrum quite similar to those observed in the *W Serpentis* stars.

The system parameters of U Cephei are much better known than those of the other members of the group. Moreover, the eclipses in U Cephei are conveniently short, so that in one or two consecutive *IUE* shifts, it is possible to scan the vicinity of the accreting star in

¹ *IUE* Guest Observer.

a search for the location and geometrical distribution of the emitting material. For these reasons, the discovery that U Cephei is a W Serpentis system has a great significance for the understanding of the phenomenon and, ultimately, for understanding the processes of mass transfer—accretion and mass loss—in the interacting close binary systems. Work on this system is greatly facilitated by previous extensive studies by Batten (1974, 1981) and Olson (1980*a, b*).

II. PARAMETERS OF THE SYSTEM

U Cephei was comprehensively analyzed by Batten (1974), and his review was recently updated (Batten 1981). Our description can therefore be brief. Batten concluded that the spectral types of the two components are probably best represented by B7 V and G8 III–IV. Several photometric studies were published since Batten's 1974 paper: Hall and Walter (1974), Markworth (1979), Broglia and Conconi (1980). The light curve is distorted, and its analysis cannot be brought to the desired level of accuracy. On the other hand, the object is bright, the eclipses are total, and the light curve is still much simpler than, for example, that of W Serpentis; therefore, all three solutions are reasonably well represented by the following parameters: Fractional radii $r(B) = 0.17$, $r(G) = 0.33$, and $i = 83^\circ$. The radial velocity curve of the B star is badly distorted, but a careful analysis by Batten (1974) combined with the recent Reticon studies of the secondary (Tomkin 1981) yields fairly trustworthy masses and dimensions, namely: masses $m(B) = 4.4 M_\odot$, $m(G) = 2.9 M_\odot$; radii $R(B) = 2.55 R_\odot$, $R(G) = 4.95 R_\odot$; and the separation of the components $A = 15.0 R_\odot$. Quite recently, Olson (1983) derived a new set of photometric parameters. Since Olson has studied U Cephei very carefully over a number of years, and, in the paper referred to, made a painstaking effort to include only data as clear of contamination as possible, his elements are likely to be preferred. However, the changes are rather small. Olson's values $r(B) = 0.183$, $r(G) = 0.324$, and $i = 85.8^\circ$ change the

dimensions to $R(B) = 2.74 R_\odot$ and $R(G) = 4.86 R_\odot$. We will accept these values in this paper.

The period of the system is variable. Fortunately, over the past three years (1979–1982) it has been constant, and the times of eclipses can be calculated with the most recent ephemeris by Olson *et al.* (1981), slightly modified as suggested in a private communication by Olson:

$$T \text{ (primary eclipse)} = \text{JD } 2438291.533 + 2^d 4930474E. \quad (1)$$

The primary eclipse starts at about phase 0.916*P*, becomes total at 0.981*P*, and the following two contacts are approximately symmetrically located at phases 0.019*P* (end of totality) and 0.084*P* (end of the partial phase).

III. THE OBSERVATIONS

Table 1 lists the observations made with the *IUE* satellite. The high-dispersion spectrum has a resolving power of 1.2×10^4 , while the low-dispersion spectrum in the short-wavelength (SWP) camera has a resolution of about 6 Å, and in the long-wavelength camera (LWR), about 8 Å (Boggess *et al.* 1978). Table 2 lists the optical scans, obtained at Lick Observatory with the Cassegrain image-tube scanner at the 3 m Shane telescope. The grating used gives nearly the same resolution, about 9 Å, as the low-dispersion *IUE* spectra. The optical spectrum between 320 and 820 nm was covered in three scans, labeled *B* (blue), *R* (red), and *I* (infrared).

The first *IUE* observations were made on 1980 May 23 in the low-dispersion mode. They showed a rather ordinary late-B spectrum with no obvious emission lines present. However, one can suspect an emission of C II at 1336 Å. The LWR spectrum is overexposed between 250 and 310 nm, and a weak emission at the Mg II doublet cannot be ruled out with certainty.

IUE observations during the total eclipse of 1981 August 2 revealed the presence of high-ionization emission lines in the region covered by the SWP camera;

TABLE 1
IUE OBSERVATIONS OF U CEPHEI

Date, UT (mid-exposure)	Image	HJD 2,440,000. +	Phase	Eclipse	Exposure (s)
1980 May 23, 1308	LWR 7830	4383.045	0.400	out of eclipse	60
1980 May 23, 1320	SWP 9083	4383.053	0.403	out of eclipse	50
1981 Aug 2, 0712	LWR 11213	4818.799	0.187	out of eclipse	40
1981 Aug 2, 0751	SWP 14624 ^a	4818.826	0.198	out of eclipse	3600 ^a
1981 Aug 4, 0809	SWP 14646	4820.838	0.0055	primary, total	600
1981 Aug 4, 0823	LWR 11237	4820.848	0.009	primary, total	600
1981 Dec 14, 0940	SWP 15746	4952.905	0.980	primary, partial	1080
1982 Aug 30, 1522	SWP 17810	5212.140	0.963	primary, partial	210
1982 Aug 30, 1530	LWR 14058	5212.146	0.965	primary, partial	150
1982 Aug 30, 1629	SWP 17811	5212.187	0.981	primary, partial	480
1982 Aug 30, 1642	LWR 14059	5212.196	0.985	primary, total	300
1982 Aug 30, 1729	SWP 17812	5212.228	0.998	primary, total	2100

^a SWP 14624 is a high-dispersion spectrum; the rest are low-dispersion spectra.

TABLE 2
OPTICAL SCANS OF U CEPHEI

Date, UT (mid-exposure)	Region	HJD 2,440,000. +	Phase	Eclipse	Observers ^a
1980 Jun 21, 1119	<i>BRI</i>	4411.969	0.002	primary, total	AK
1981 Aug 8, 1028	<i>BRI</i>	4824.935	0.649	out of eclipse	PW
1981 Oct 13, 0735	<i>BRI</i>	4890.818	0.075	primary, partial	PW
1982 Aug 5, 1208	<i>BRI</i>	5187.004	0.880	out of eclipse	PW
1982 Aug 28, 0654	<i>BRI</i>	5209.787	0.019	primary, total	KW
1982 Aug 30, 1040	<i>BRI</i>	5211.944	0.884	out of eclipse	KW
1982 Oct 1, 0833	<i>BRI</i>	5243.858	0.681	out of eclipse	PW
1982 Oct 2, 0430	<i>BRI</i>	5244.689	0.014	primary, total	PW

^a Observers: A = L. H. Aller; K = C. D. Keyes; P = M. J. Plavec; W = J. L. Weiland.

also, the Mg II doublet was very clearly in emission in the LWR spectrum. The emissions were then observed again on 1981 December 14 at a phase just preceding the totality. Then, on 1982 August 30, a sequence of five *IUE* spectra was taken through the primary eclipse in an attempt to locate the region in which the emissions are formed. During the observations, the brightness of the star changed so rapidly that it was difficult to make adequate exposures. Thus, most of the spectrum LWR 14058 is overexposed, and so is the long end of the spectrum SWP 17810.

Scanner observations of U Cephei at Lick Observatory in 1980 and 1981 provided some information about the optical flux distribution, but really useful data were not obtained until 1982.

IV. FLUX DISTRIBUTION IN THE CONTINUOUS SPECTRUM

The reddening of U Cephei is small, as it should be for a nearby star rather far from the galactic plane ($b = +17^\circ$). There is no measurable 220 nm interstellar bump. Olson (1980a) adopted $E(b - y) = 0.05$ mag, so that $E(B - V) = 0.03$ or 0.04 mag. Eggen (1963) concluded that $E(B - V) = 0.03$ mag. In our program FITMEX, which matches the observed flux distribution with the Kurucz (1979) model atmospheres, the color excess is one of the unknowns to be determined by the least squares solution; and our solutions converged to $E(B - V) = 0.03$ mag. The Kurucz model fitting described in this section was done mostly by Jan J. Dobias as part of our systematic survey of the Algol systems, to be published separately (J. J. Dobias and M. J. Plavec, in preparation; see also Plavec and Dobias 1982).

The out-of-eclipse observations were used to determine the spectral energy distribution of the B star. The spectral region 118–600 nm was represented by 192 discrete wavelengths for which Kurucz tabulated the model fluxes. Although our Lick scans go well beyond 600 nm, those points were not used since those fluxes are already perceptibly contaminated by the secondary component. There exists a gap between the Lick data and the *IUE* data in the region 310–330 nm, where neither instrument provides reliable coverage. In addition, we have established that the fluxes between

260–310 nm (*IUE* data) and again between 330–375 nm (Lick data) were contaminated by an additional source, almost certainly by bound-free optically thin radiation of circumstellar hydrogen.

The out-of-eclipse observations with the *IUE* and with the Lick scanner were not made at the same phase; in fact, the dates differ by more than 1 yr. Since serious phase-dependent and epoch-dependent differences may be present, separate fits were made at first for the *IUE* data (31 points between 118–200 nm) and for the Lick scans (71 points between 375–600 nm). The optical flux distribution is best fitted by $T_{\text{eff}} = 11,000$ K, $\log g = 3.75$, $E(B - V) = 0.05$ mag.

The *IUE* data presented a special problem. The low-dispersion spectrum of U Cephei shows more numerous and deeper depressions than any standard or normal star of a similar spectral type. Hence its absorption lines in the far-ultraviolet must be both deeper and more numerous. This is most likely due to the existence of a circumstellar envelope to be discussed in § VII. Thus, as a first step, wavelengths were selected in the high-dispersion spectrum of U Cephei for which the observed flux (the so-called FN numbers) is a local maximum. These points are believed to be much closer to the normal quasicontinuum than any other points in the spectrum of U Cephei. Figure 1 shows that these selected points agree very well with the highest points in the low-dispersion SWP spectrum. These points, then, were matched by various Kurucz model atmospheres. The best fitting interpolated atmospheric model is that for $T_{\text{eff}} = 11,250$ K, $3.75 \leq \log g \leq 4.0$, and $0.00 \leq E(B - V) \leq 0.04$ mag. The results for the SWP spectral range and for the optical spectral range are so close that one fit to the combined spectrum is justified. This fit is $T_{\text{eff}} = 11,250$ K, $3.75 \leq \log g \leq 4.0$, and $E(B - V) = 0.03$ mag. Therefore, the hotter star in U Cephei is of spectral type B8 V.

The LWR spectrum, covering the wavelength interval between about 190 and 310 nm, was not used in the fitting. For this spectral region, we do not have a corresponding high-dispersion spectrum, so the corrections for additional line-blocking could not be made. Moreover, as Figure 2 shows, from about 260 nm longward the flux distribution cannot be matched by any Kurucz

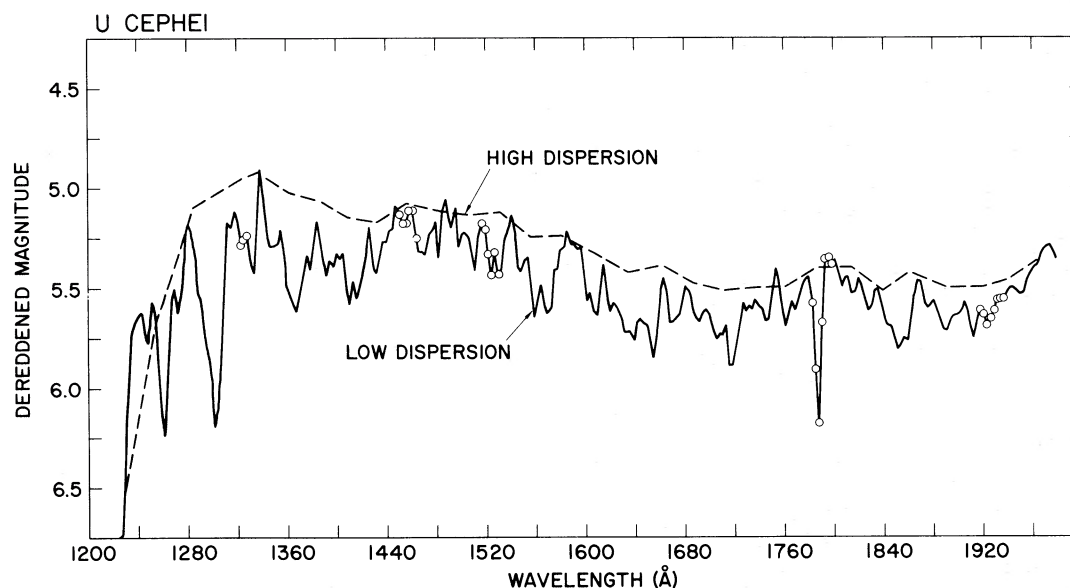


FIG. 1.—The far-ultraviolet flux of U Cephei outside eclipse as recorded at low dispersion (*solid line*) is strongly depressed by line blocking. The true continuum is more likely represented by several apparently line-free segments in the high-dispersion spectrum, connected here by a broken line. The open circles in the low-dispersion spectrum mark reseau (at which the fluxes are unreliable).

atmospheric model, and an additional radiating source must be assumed, most likely a circumstellar optically thin hydrogen cloud.

As explained elsewhere (Plavec, Weiland, and Koch 1982), from the difference between the observed (but dereddened) monochromatic magnitudes m_λ and the magnitudes $M_\lambda = -2.5 \log F_\lambda$ derived from the fluxes tabulated by Kurucz, we obtain the ratio R/d of the

stellar radius to the star's distance:

$$m_\lambda - M_\lambda = 33.321 - 5 \log (R/d), \quad (2)$$

where $R = R(B)$ is in solar radii, and d is in kpc. The least squares solution gave 2.77 ± 0.05 mag for the shift; therefore, the most probable value of R/d is 12.88. If $R(B) = 2.55 R_\odot$ as derived above, the distance comes out to be 198 pc. Adopting Olson's value of $R(B) = 2.74 R_\odot$

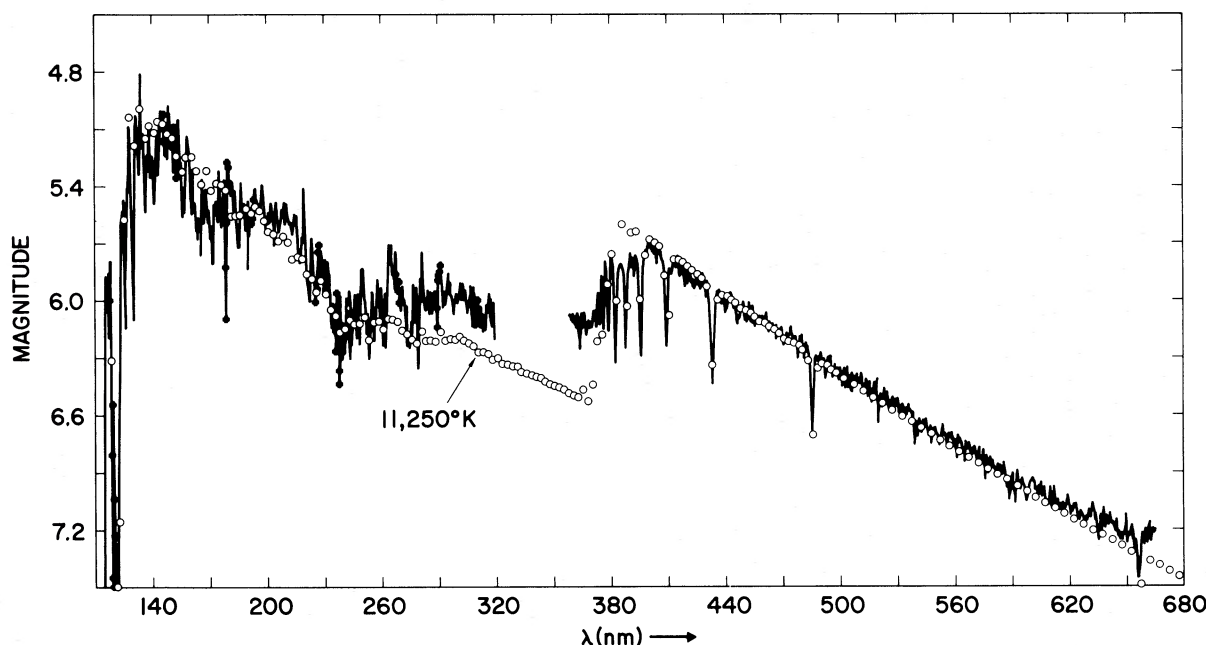


FIG. 2.—The overall best fit to the observed flux distribution of U Cephei (*solid line*) is obtained from an interpolated Kurucz atmospheric model with $T_{\text{eff}} = 11,250$ K, $\log g = 4.0$ (*open circles*). Filled circles are reseau.

changes the distance to $d = 213$ pc. After approximate weighting of all the solutions, I decided to adopt $d = 208$ pc. The resulting distance is in excellent agreement with Batten's (1974) conclusion that the distance must be about 200 pc. We will adopt a distance modulus of 6.59 mag. It is difficult to estimate how accurate our distance determination is; we hope that it lies within about 10% of the correct value.

The V_0 (dereddened) magnitude of the B component derived from the Lick scan is 6.85 mag, after correction has been made for the contribution of the secondary star. The resulting $M_v(B) = +0.26$ mag. This agrees with Blaauw's luminosity calibration (Blaauw 1963): he gives $+0.1$ mag for B8 V. Earlier calibrations tended to assign higher absolute magnitudes to B8 V stars, but this was due to a too-high effective temperature adopted for them then. It is preferable to use the critical compilation of empirical data collected by Popper (1980) from his binary star studies. For spectral type B8 V, all values $0.0 \text{ mag} \leq M_v \leq 1.0 \text{ mag}$ are acceptable, according to Popper's data.

Batten (1974) discussed at length the existing uncertainty as to the spectral type and absolute visual magnitude of the B star, which was variously classified in the range B6–A0. Our result may remove part of the uncertainty. The effective temperature derived is lower than most of the other estimates, which were, however, based on much less complete data or estimated from the line spectrum. The best spectral type determination by Olson (1980a) was based on the Strömgren photometry and led to $T_{\text{eff}} = 13,000$ K, $\log g = 4.0$, and $E(B-V) = 0.03$ mag. The flux from the B star may be variable and

depend on the orbital phase and on the instantaneous state of activity of the system, since Olson found that optically thick transient disks or shells often act as an additional source of radiation. No such source was detected in our data (except for the bound-free optically thin hydrogen radiation). Our data do not allow an effective temperature as high as 13,000 K, as can be seen from Figure 3. A comparison with our preferred fit in Figure 2 shows that the slope of the Paschen continuum varies rather slowly with temperature, making its determination rather uncertain. The addition of the *IUE* spectrum, which was not accessible to Olson, considerably constrains the acceptable models and definitely demands an effective temperature near 11,000 K. But our optical data are also better represented by a temperature this low; if anything, an even lower temperature appears suitable. We conclude that at the time of our observations, the spectral type of the B star was B8 V.

Our best overall fit, shown in Figure 2, is not as satisfactory as what we have found in other Algol primaries. However, the derivations are local rather than systematic. The ultraviolet absorption lines in U Cephei are deeper than in single stars of the same type; no doubt some circumstellar material is present even when the light curve appears unusually clean. On the contrary, there is excess observed flux around 210 nm, perhaps related to the peculiar bulge observed to be very conspicuous in the W Serpentis stars in that region (Plavec 1982a; Plavec, Weiland, and Dobias 1982). Significant also seems to be the flux deficiency between 380 and 400 nm.

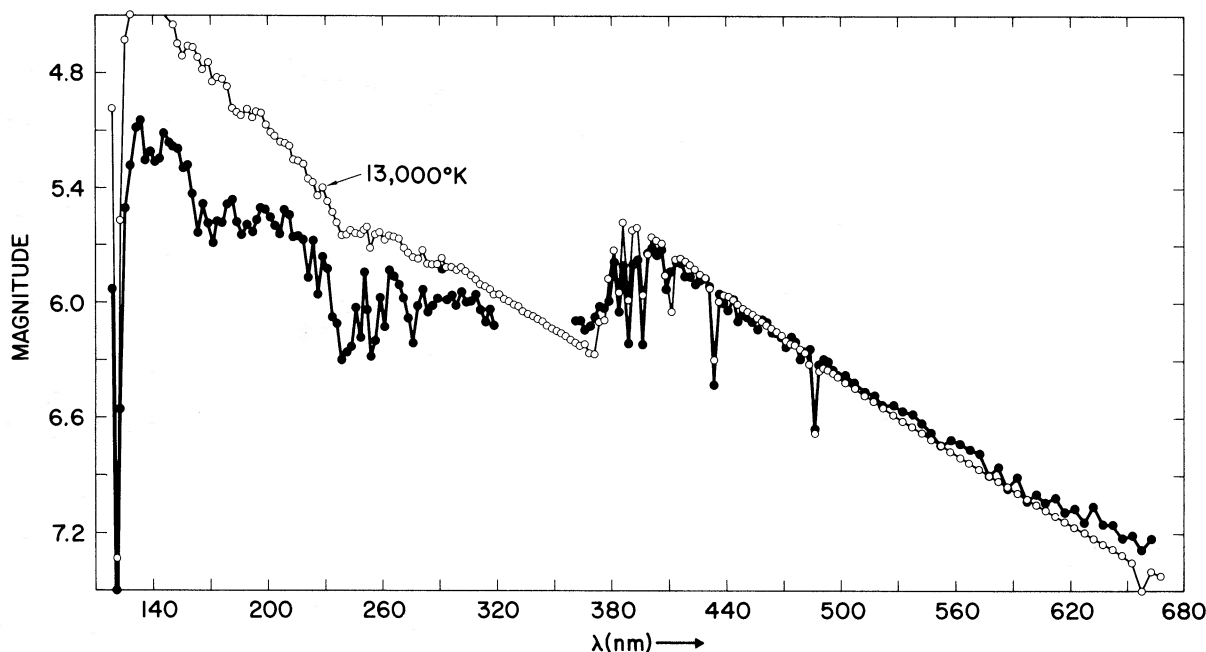


FIG. 3.—An effective temperature of 13,000 K (open circles), often adopted for U Cephei, provides a fairly good match in the optical spectral region but fails completely to match the ultraviolet part of the observed flux distribution, simplified here into the “Kurucz wavelength points” (filled circles).

The Balmer jump, as Figure 2 shows, is definitely shallower than for any model atmosphere within the acceptable temperature range. Since there is also a systematic flux excess between 260 and 310 nm, we conclude that optically thin bound-free radiation of hydrogen contributed to the light at that particular phase at both epochs. We plan new observations in order to study this circumstellar radiation in detail.

It is obvious that the Balmer jump cannot be used to measure the surface gravity, which is therefore rather weakly determined. The adopted mass and radius lead to $\log g = 4.26$, which is at the edge of our error box. The B star is a main-sequence star. We are unable to refine our analysis so as to take into account the flattening of the B star due to its rapid rotation.

A pure spectrum of the G component was observed optically during the total eclipse on 1980 June 21. The overall optical flux distribution is very well matched by the G8 III standard star ξ Herculis over the whole interval 320–680 nm, with the exception of a narrow segment near 390 nm, where the standard star has deeper absorption lines than the late-type component of U Cephei (Fig. 4).

There is only a weak indication of an excess flux shortward of the Balmer limit; the hydrogen cloud, whose radiation was seen in the out-of-eclipse scans discussed above, was either largely absent or eclipsed by the cool star. From our scan, we obtained $V_0(G) = 9.25$ mag, leading to $M_v(G) = +2.8$ mag. Blaauw's calibration (1963) gives $M_v = +3.0$ mag for G8 IV and $+0.3$ mag for G8 III. Thus, the star is a G8 IV subgiant. We do not have any scan of a G8 IV standard

star and do not know if substantial differences could be detected between luminosity classes III and IV. Broglia and Conconi (1980) found a somewhat brighter magnitude from their *UBV* photometry, $V = 9.06$ mag. The difference may be real, because the star is almost certainly variable, but it does not alter the luminosity classification. As to the temperature constraints, both G5 III and K1 III standards gave significantly poorer fits than G8 III.

V. THE EMISSION-LINE SPECTRUM

The emission lines are visible only during the total primary eclipse, and at phases closely adjacent, as the 1982 August 30 observations show. At phase 0.963, two-thirds through the partial eclipse, the spectrum is still essentially that of the primary star, although weakened by about a factor of 3, and emission lines are at most marginally visible. Two spectra centered just at the beginning of totality (SWP 15746, taken on 1981 December 14, and SWP 17811, taken on 1982 August 30) are already dominated by emission lines, although a definite continuum is still present. The two spectra obtained during totality are pure emission-line spectra, with a very low and flat continuum. Figure 5 shows the 1981 August 4 totality spectrum.

The SWP 17811 spectrum is unusually rich in emission lines (Fig. 6), although at places it is hard to distinguish between noise and real lines. Also, the presence of a stellar continuum with many absorption lines makes it less easy to determine the line fluxes. It has therefore been decided to concentrate first on the totality spectra. Among them, the 1981 August 4

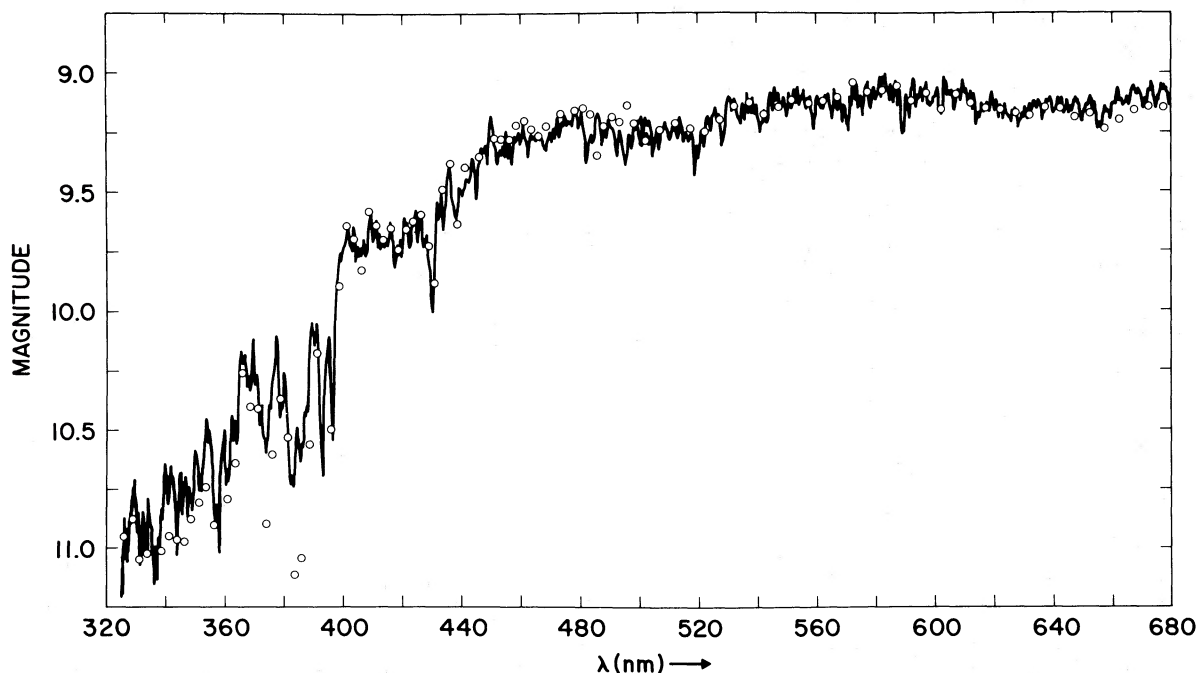


FIG. 4.—The flux distribution in the optical region observed during the total primary eclipse (solid line). It is well matched by the standard star ξ Herculis (G8 III) (circles).

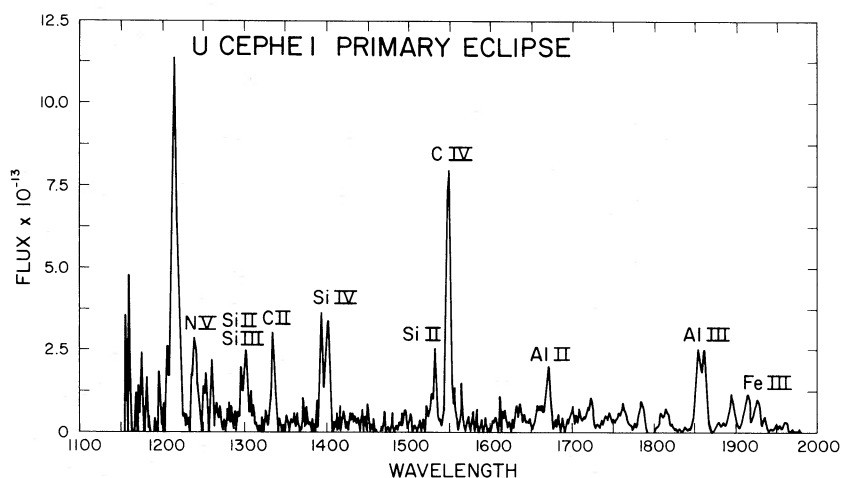


FIG. 5.—The emission-line spectrum observed by the SWP camera during the total eclipse on 1981 August 4.

spectrum (SWP 14646) has somewhat stronger and better defined emission lines. This spectrum has therefore been selected as the basis for the emission-line list in Table 3. The 1981 December 14 spectrum (SWP 15746) shows a very similar pattern of lines and is smoother than SWP 17811, so line fluxes are listed in Table 3 also for this spectrum. We will first discuss the totality spectrum, and we will take up the comparison with the out-of-totality spectra in the subsequent section.

Table 3 lists the fluxes of the emission lines in the 1981 totality spectrum, $F(T)$, and in the 1981 pretotally spectrum, $F(P)$. These are the observed fluxes, but corrected for interstellar reddening. In both columns, the unit is 10^{-12} ergs cm^{-2} s^{-1} . Where two or three lines are blended (although their peaks can be separated), the letter *b* is in the last column. The fourth column of

Table 3 contains the actual powers emitted in the emission lines, expressed in units of $10^{-4} L_{\odot}$, for the 1981 totality spectrum (SWP 14646). The following columns then identify the contributors to the line; in addition to the multiplet number taken from Moore (1950), we list the laboratory wavelength and the excitation potential of the lower (LEP) and upper level (UEP), respectively. The subsequent columns refer to the absorption spectrum and will be explained in § VII.

The identification of the contributing lines was straightforward with a few exceptions. The blended strong line at 1403 Å may also have the intercombination multiplet of O IV (1) as a contributor, in addition to Si IV. At low dispersion, it is impossible to decide if it is present. The Si III] (1) 1892 Å line is blended with the Fe III (34) line at 1895.5 Å, and since the

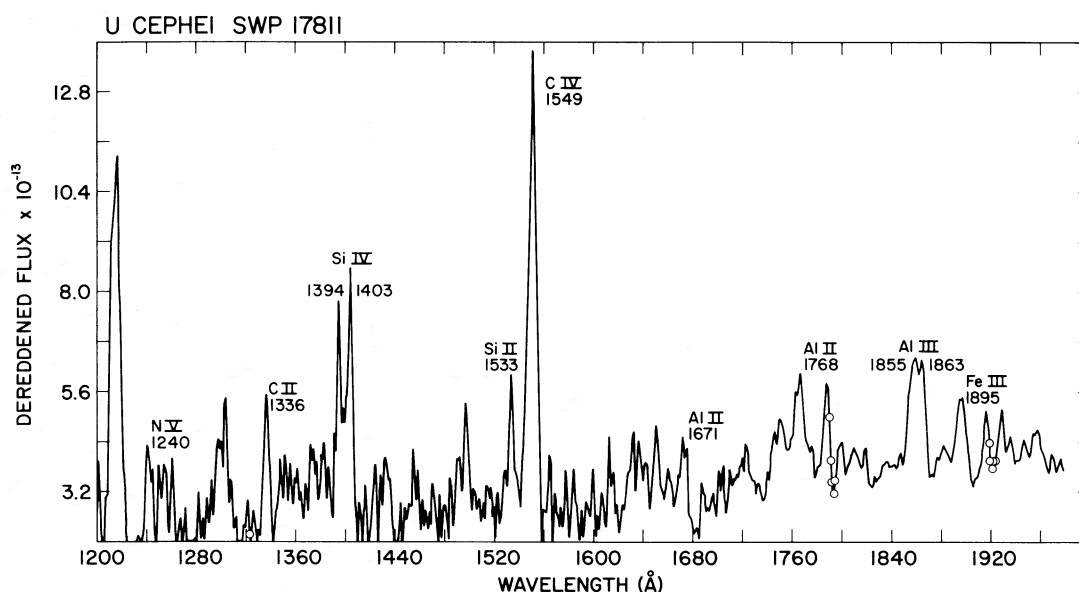


FIG. 6.—Spectrum SWP 17811 observed on 1982 August 30 at phase 0.981

TABLE 3
 PROMINENT EMISSION LINES OBSERVED IN U CEPHEI

λ_{obs} (Å)	$F(T)$ 1981 (10^{-12} ergs $\text{cm}^{-2} \text{s}^{-1}$)	$F(P)$ 1981 (10^{-12} ergs $\text{cm}^{-2} \text{s}^{-1}$)	$P(T)$ 1981 ($10^{-4} L_{\odot}$)	ID	λ_{lab} (Å)	LEP (eV)	UEP (eV)	ABSORPTIONS		NOTES
								EW (Å)	r_c	
1240.....	1.31	1.76	17.2	N v (1)	1238.8	0.0	10.0	0.8	0.50	
				N v (1)	1242.8	0.0	10.0	0.6	0.52	
1253.....	0.59	0.46	7.7	S II (1)	1250.5	0.0	9.8	1.6	0.28	
				S II (1)	1253.8	0.0	9.8	0.8	0.46	
1261.....	0.54	...	7.1	Si II (4)	1260.4	0.0	9.8	2.9	0.15	
				Si II (4)	1264.7	0.0	9.8	3.3	0.12	
				Si II (4)	1265.0	0.0	9.8	...	...	1
1296.....	0.56	...	7.4	Si III (4)	1296.7	6.6	16.1	1.1	0.40	2, b
				Si III (4)	1298.9	6.6	16.1	1.2	0.30	
				O I (2)	1302.2	0.0	9.5	...	...	
1302.....	1.14	0.93	15.0	Si II (3)	1304.4	0.0	9.5	...	...	
				O I (2)	1304.9	0.0	9.5	...	...	1, 3, b
1308.....	0.20	...	2.7	Si II (3)	1309.3	0.0	9.5	...	...	
1335.....	1.65	2.62	21.6	C II (1)	1334.5	0.0	9.3	2.7	0.13	
				C II (1)	1335.7	0.0	9.3	...	...	1
1394.....	1.40	2.67	18.4	Si IV (1)	1393.8	0.0	8.9	1.8	0.28	4
1403.....	1.95	2.87	25.6	Si IV (1)	1402.8	0.0	8.8	1.3	0.39	
1533.....	1.26	1.30	16.6	Si II (2)	1526.7	0.0	8.1	1.3	0.31	4
				Si II (2)	1533.4	0.0	8.1	1.5	0.33	
1549.....	5.11	7.30	67.2	C IV (1)	1548.2	0.0	8.0	0.9	0.44	
				C IV (1)	1550.8	0.0	8.0	1.0	0.39	
1565.....	0.30	...	4.0	Si II (10.02)	1563.8	6.9	14.8	...	...	5
1671.....	1.05	1.33	13.8	Al II (2)	1670.8	0.0	7.4	1.4	0.32	4
1723.....	0.43	...	5.7	Al II (6)	1719.5	4.6	11.8	...	...	
				Al II (6)	1721.3	4.6	11.8	...	...	
				Al II (6)	1725.0	4.6	11.8	...	...	
1763.....	0.50	...	6.5	Al II (5)	1763.9	4.6	11.6	...	...	
				Al II (5)	1767.6	4.6	11.6	...	...	
1785.....	0.45	...	5.9	Ni II (5)	1788.5	0.2	7.1	...	...	6
1809.....	0.18	...	2.4	Si II (1)	1808.0	0.0	6.9	0.9	0.59	
1816.....	0.38	...	5.0	Si II (1)	1816.9	0.0	6.9	1.7	0.47	7
1854.....	1.70	1.36	22.2	Al III (1)	1854.7	0.0	6.6	1.3	0.40	b
1862.....	1.24	1.81	16.3	Al III (1)	1862.8	0.0	6.6	1.8	0.38	b
1895.....	0.77	0.85	10.2	Si III (1)	1892.0	0.0	6.6	...	...	
				Fe III (34)	1895.5	3.7	10.2	1.3	0.54	
1915.....	0.68	0.56	8.9	Fe III (34)	1914.1	3.7	10.2	1.1	0.56	
1926.....	0.66	0.34	8.6	Fe III (34)	1926.1	3.7	10.1	1.1	0.58	
2799.....	2.85	...	3.35	Mg II (1)	2795.5	0.0	4.4	...	...	8
				Mg II (1)	2802.7	0.0	4.4	...	...	

NOTES.—(1) Absorption line blended with the preceding line. (2) Absorption line distorted (edge of a spectral order). (3) Absorption line 1306.0 Å of O I (2) not seen. (4) Absorption line has an extended wing shortward. (5) Blended with Si II (10.02) 1564.1 Å. These lines should be rather weak, and another identification was considered, mainly with Fe II. However, no other Fe II emission lines were found with certainty. (6) If the identification with Ni II (5) is correct, one should expect other Ni II (5) lines, in particular 1741 and 1748 Å (referee's comment). They are probably indeed present but not conspicuous. (7) Absorption of Si II (1) 1817.4 Å not seen. (8) Partial phase was not observed in LWR; no high-dispersion LWR spectrum was available. (b) The line is not fully resolved from its neighbor.

blended line is centered near 1895 Å, the latter contributor is dominant. Thus from among the intercombination multiplets falling into the region covered by the SWP camera, only O III] at 1663 Å, N IV 1487 Å, N III] at 1750 Å, and C III] at 1909 Å would be distinguishable if present with sufficient intensity. They have not been found. As they cover a fairly wide range of both electron temperature and critical electron density, a representative electron density of at least $5 \times 10^{11} \text{ cm}^{-3}$ is indicated. This would be in agreement with our previous findings in SX Cassiopeiae (Plavec, Weiland, and Koch 1982). While there remains, in both cases, a degree of uncertainty because of the low spectral resolution, our detailed study of the high-dispersion spectra of β Lyrae confirms the absence of all intercombination lines with the possible exception of Si III] (1) 1892 Å. That line was used for a quantitative evaluation, and it was concluded that in β Lyrae, $n_e \geq 5 \times 10^{12} \text{ cm}^{-3}$. We do not know if this result is valid for all the Serpents, but the similarity of the emission spectra certainly makes it possible.

It is instructive to compare Table 3 with the similar Table 4 in our article on SX Cas. The same emission lines are observed in both stars, with very few and insignificant exceptions, as follows. The P III (1) lines at 1344.3 and 1344.9 Å can be distinguished as a probably real but weak double emission in SX Cas, but in U Cep the fluxes do not exceed the noise level. The same is true about the Ni II (7) zero-volt lines at 1455 and 1500 Å. Another multiplet of nickel, Ni II (5) at 1741.6 and 1748.3 Å, gives a little stronger emission in SX Cas and appears to correspond to a somewhat stronger feature also in U Cep, but it still remains within the noise level, so it is not listed in Table 3. A number of features at about the level of this line appear in the spectrum of U Cep, and some of them are probably real and identifiable lines; however, we decided to wait for additional spectra before we list them with certainty as real emission lines. The emission lines in U Cep appear to be somewhat narrower than in SX Cas, so that more lines can be resolved. The better resolution in U Cep partly explains different appearances of several blends. Thus, the Al III (1) doublet at 1854.7 and 1862.8 Å is clearly resolved in two peaks, while the resolution is only partial in SX Cas, and the line was listed by us as a broad line centered at 1860 Å, although its double character is evident in SX Cas, too. The broad blend near 1300 Å deserves special attention. In U Cep we find a fairly strong peak at 1302 Å, a lower one at 1296 Å, and two still smaller ones at 1308 and 1311 Å. In SX Cas, there is a fairly high peak at 1300 Å, another only slightly lower at 1305 Å, and a much weaker one at 1311 Å. This region includes the subordinate multiplet Si III (4) with six lines between 1294 and 1305 Å, and zero-volt multiplets of Si II (3) (lines at 1304 and 1309 Å) and of O I (2) (lines at 1302, 1305, and 1306 Å). The high-dispersion spectra of β Lyrae show that the dominant contributor is Si III (4); Si II (3) lines are present but weaker, and O I (2) lines are missing. This distribution is degraded, in the low-dispersion spectra of

β Lyrae, into one high peak at 1299 Å and a well-separated, much lower peak at 1305 Å. Thus the peak at 1299 Å is to be identified solely with the Si III (4) lines, while the peak at 1305 Å corresponds to Si II (3), although the Si III lines at 1301 and 1303 Å probably also contribute. We conclude that in SX Cas, Si II (3) is relatively stronger compared with Si III (4), but a contribution from O I (2) cannot be excluded. The same appears to apply to U Cep, where the whole blend is somewhat better resolved.

Since the emission spectra of SX Cas and U Cep are qualitatively so similar, let us compare the line intensities. From comparisons of high- and low-dispersion spectra of β Lyrae, we know that the low dispersion badly degrades the line profiles, so badly indeed that the pronounced P Cygni character seen in high dispersion is completely obliterated. Having nothing better than the low-dispersion spectra of U Cep and SX Cas, we must attempt to compare at least those lines that are not severe blends of lines coming from different species.

Table 4 compares the line intensities within each spectrum, and between the spectra of SX Cas and U Cep. The reference line was chosen as follows. The last column contains the ratio of the powers emitted in SX Cas and in U Cep. The mean of these ratios is 138. Several lines have intensity ratios close to this number. Among them, Al II (2) at 1671 Å appears to be most suitable, since it is a single, well-defined line. Therefore, this line was selected as a reference line in both spectra, and the corresponding internal ratios are also listed in Table 4.

The most striking observation is that the emission lines in SX Cas are much stronger than those in U Cep by an average factor of more than 100. This explains their different behavior outside eclipse. In SX Cas, the (observed but dereddened) peak flux in the lines of N V, Si IV, and C IV reaches about $4 \times 10^{-12} \text{ ergs cm}^{-2} \text{ s}^{-1} \text{ Å}^{-1}$, and outside the eclipse, the adjacent continuum flux rises to just about the same level. Thus, even in the combined flux, the emission lines would remain easily visible even if they were not strengthened outside eclipse. But most of these lines are strengthened. Evidently, the space immediately surrounding the primary component in SX Cas, and eclipsed with it in totality, contributes considerably to the emission in SX Cas. In U Cep, the strongest line, that of C IV, reaches a peak flux of $7 \times 10^{-13} \text{ ergs cm}^{-2} \text{ s}^{-1} \text{ Å}^{-1}$, but the stellar continuum out of eclipse is near $3 \times 10^{-11} \text{ ergs cm}^{-2} \text{ s}^{-1} \text{ Å}^{-1}$, i.e., more than 40 times higher. A qualitatively similar situation exists in the optical region. While in SX Cas the first six of the Balmer lines probably always appear in emission (Struve 1944), in U Cep only H α has ever been seen in emission outside the eclipse, and this only during a conspicuous outburst (Batten *et al.* 1975).

If we add together all the power emitted in the emission lines included in our lists, and contained between the wavelengths 123–195 nm, we find a value of $2.8 L_{\odot}$ in SX Cas and $0.026 L_{\odot}$ in U Cep. The integrated flux emitted within the same wavelength range in the continuum (and in the unrecognized emission lines) is $3.5 L_{\odot}$ in SX Cas and $0.020 L_{\odot}$ in U Cep. Thus, the

TABLE 4
COMPARISON OF U CEPHEI WITH SX CASSIOPEIAE

LINE		POWERS, $P(\lambda)$ ($10^{-3} L_{\odot}$)		RATIOS OF POWERS		
				$P(\lambda)/P(1671)$		$P(SX)/P(U)$
Ion	$\lambda(\text{\AA})$	U Cep	SX Cas	U Cep	SX Cas	
N v	1240	1.90	316	1.7	3.1	166
Si II (4)	1264	0.48	177	0.4	1.75	369
Si III (4)	1298	0.53	165	0.5	1.6	311
Si II (3)	1303	1.15	155	1.0	1.5	135
C II (1)	1336	1.56	181	1.4	1.8	116
Si IV (1)	1394	1.43	186	1.3	1.85	130
Si IV (1)	1403	2.21	174	2.0	1.7	79
Si II (2)	1533	1.24	101	1.1	0.9	81
C IV (1)	1550	5.69	174	5.1	1.7	31
Al II (2)	1671	1.11	105	1.0	1.0	95
Al III (1)	1854	1.87	94	1.7	0.9	50
Al III (1)	1862	1.26	76	1.1	0.75	60
Fe III ^a (34)	1895	0.84	145	0.8	1.4	173
Fe III (34)	1915	0.74	101	0.7	1.00	136
Fe III (34)	1926	0.73	105	0.7	1.0	144

^a Possibly blended with Si III] 1892 Å.

recognized emission lines contribute very significantly, about 50%, to the total energy output in the region 123–195 nm during the total eclipse. Also, it seems that continuous radiation in that region is powered by the same mechanism, since the flux ratios are roughly the same for both the strong emission lines as well as for the continuum.

Why is the FUV radiation produced in SX Cas so much stronger than in U Cephei? One obvious answer is that the emitting volume may be much larger in SX Cas, since that system is much larger: the separation of the components is probably about $87 R_{\odot}$, while in U Cep it is only $15 R_{\odot}$. If the emitting volume scales as the volumes of the critical Roche lobes of the primary components, then (because the electron density is probably nearly the same) the emission measures should be in a ratio of about 300. If, alternatively, the lines are optically thick or the emitting region is disk-shaped, one would expect scaling only with the square of the radius of the critical Roche lobe, and the ratio would be about 50. However, the projected area actually observed at totality is only the area between the radius of the Roche lobe, R_1^* , and the radius of the secondary component, R_2 . These areas give a ratio of 100, very close to the observed average ratio of the emission-line intensities. This coincidence may be fortuitous, though, since the difference in emitted powers may be due to a larger energy input into the emitting process in SX Cas, due probably to a higher rate of mass influx dM/dt . Without understanding the mechanism by which the emission is generated, we will not be able to decide. But it is clear that the primary components are not involved, since their powers, emitted in the same spectral region in full light, are 30.9 and $19.8 L_{\odot}$, respectively, for SX Cas and U Cep.

The most striking difference between the two stars in

the emission-line intensities is in the behavior of C IV (1). This line is by far the strongest in U Cep, while in SX Cas it is no stronger than each of the Si IV (1) doublet lines, or the blended intensity of the multiplets of Si II (4), Si III (4), or C II (1). The equal strength of the C II and C IV resonance lines in SX Cas is of particular importance. The strongest emission line in SX Cas is the blended doublet of N v (1). In U Cep, the aluminum lines are stronger—in particular, Al III (1) at 1855 Å, which is as strong as N v (1). If we refer the line intensities to an average representing the star and compare those, we find that C II (1) appears to be roughly of the same relative strength in both stars, while C IV (1) is substantially enhanced in U Cep. These differences in relative line intensities are probably due to different temperatures in the emitting layers of the two stars. It is also possible, though, that the C/N ratio may be affected by the presence of the CNO processed material in the circumstellar matter.

A very important feature that the spectra of SX Cas and U Cep have in common is the absence of the emission line of He II 1640 Å. While this line is quite prominent, e.g., in the spectra of the symbiotic stars, it is conspicuously absent in the spectra of all the W Serpentis stars. In AG Pegasi (Plavec 1982b) and in similar symbiotic stars, the rising FUV continuum indicates the presence of a hot ($T_{\text{eff}} \approx 10^5$ K) subdwarf which is the source of the ionizing photons. In other symbiotic stars, the subdwarf may (possibly!) be replaced by an accretion disk with a hot transition zone between the disk and the central star. In SX Cas and U Cep we do not observe any similar hot continuum. Should such a continuum be strong only at even shorter wavelengths, one would expect to observe X-rays coming from the system. An attempt to detect the X-rays from the W Serpentis stars with *Einstein* failed (Guinan, Koch,

and Plavec 1983). Only a weak X-ray source was detected in the noneclipsing object V644 Monocerotis (HD 51480), which may be an interacting binary akin to the Serpentids. In the case of the eclipsing Serpentids like SX Cas or W Ser, one could argue that the X-rays are absorbed either in a thick disk surrounding the accreting star or in the interstellar medium, since both stars lie very near the galactic plane and are rather heavily obscured. This argument cannot be used for U Cephei, which is a fairly bright nearby star, almost unreddened, and no thick ring surrounds it (perhaps with the exception of the periods of strong H α activity).

U Cephei was observed with *Einstein* by White and Marshall (1982), who detected a weak but definite X-ray flux, from which they derived an X-ray luminosity $L_x = 7.6 \times 10^{30}$ ergs s $^{-1}$. The distance to U Cep they assumed, 230 pc, is almost certainly somewhat large. The X-ray luminosity obtained by adopting our distance of 208 pc is 6×10^{30} ergs s $^{-1} = 0.0015 L_\odot$. White and Marshall argue, rather convincingly, that the X-ray flux coming from U Cep and similar Algol-type eclipsing binaries originates in the corona of the late-type component and is probably determined by their velocity of axial rotation. There seems to be no link to mass transfer. In any case, the above X-ray luminosity is too weak to be responsible for the photoionization of the observed FUV emission lines: it is 17 times weaker than the power emitted in the strong emission lines alone. The absence of any other powerful photoionizing source, contrasted with the presence of strong emission lines of C IV and N V, suggests that the ionization is predominantly collisional.

VI. LOCATION OF THE LINE-EMITTING REGION

In order to understand the nature of the line-emitting mechanism, it would be very important to localize the emitting region(s). A first attempt toward this goal was made by running a sequence of spectra through the

eclipse on 1982 August 30. The rather long recharge times between two successive spectra taken with the *IUE* permitted us to obtain only three SWP spectra and two LWR spectra—much less than desired for our purpose. Combining spectra obtained during different eclipses cannot give the desired information, since a large non-periodic variation is almost certainly present. This is seen from Table 5, which compares the most prominent emission lines in two pairs of spectra: the two totality spectra and two near-totality spectra. The most convincing argument in favor of nonperiodic changes in emission-line intensities comes from a comparison of the two pretotality spectra, SWP 15746 (1981 December 14) and SWP 17811 (1982 August 30). The phases of mid-exposures are nearly identical: 0.9796 and 0.9814, respectively. Figure 7 shows that the two exposures overlap and coincide very well in phase. Also, the shapes and levels of the underlying continua agree closely. Yet the 1982 spectrum is much richer in emission lines (Fig. 6), and these are also stronger (Table 5), with few exceptions.

Thus, when attempting to localize the line-emitting regions, we can safely compare only the two spectra taken during the 1982 August 30 eclipse: for these, the differences in line strengths must be due to the changing geometry of the visible regions as they are eclipsed by the cool star. And there are perceptible differences in line strengths between these two spectra. In the pretotality spectrum (marked *P*), the emissions have, on the average, higher peak intensities by 97% and higher total fluxes by 115%. (Needless to say, the differences in the continuum levels were taken into account.) This result is highly significant in itself. One of the possible explanations of the presence of the far-ultraviolet emission lines in the Serpentids has been (and in the opinion of some colleagues still is) that they originate in the chromosphere (or the transition layer between the chromosphere and the corona) of the late-type component. Several arguments were presented

TABLE 5
COMPARISON OF THE 1981 AND 1982 DATA

LINE		PEAK FLUXES, F_{peak} (10^{-13} ergs cm $^{-2}$ s $^{-1}$)				TOTAL LINE FLUXES, F (10^{-13} ergs cm $^{-2}$ s $^{-1}$)			
Ion	$\lambda(\text{\AA})$	82T	81T	82P	81P	82T	81T	82P	81P
N V	1240	1.63	2.70	2.68	2.83	11.0	13.1	18.7	17.6
C II (1)	1336	1.82	2.91	3.58	3.62	9.0	16.5	20.8	26.2
Si IV (1)	1394	2.67	3.47	5.80	4.42	12.8	13.9	28.8	26.7
Si IV (1)	1403	2.94	3.25	6.67	5.21	13.6	19.5	35.3	28.7
Si II (2)	1533	1.18	2.38	3.49	1.83	6.5	9.1	19.3	13.0
C IV (1)	1550	6.04	7.62	11.90	8.74	38.8	51.1	83.6	73.0
Al II (2)	1671	1.38	1.96	1.58	1.66	8.1	10.5	11.7	13.3
Al III (1)	1854	1.68	2.41	2.86	2.21	11.9	16.9	25.9	13.6
Al III (1)	1862	1.49	2.44	2.76	1.09	10.0	12.4	18.0	18.1
Fe III (34)	1895	0.85	1.11	1.74	1.19	8.0	7.7	18.7	8.5
Fe III (34)	1915	0.82	1.08	1.38	0.97	5.8	6.8	8.2	5.6
Fe III (34)	1926	0.59	0.94	1.35	0.53	3.4	6.6	8.3	3.3

NOTE.—82T = SWP 17812, phase = 0.998; 82P = SWP 17811, phase = 0.981; 81T = SWP 14646, phase = 0.0055; 81P = SWP 15746, phase = 0.980.

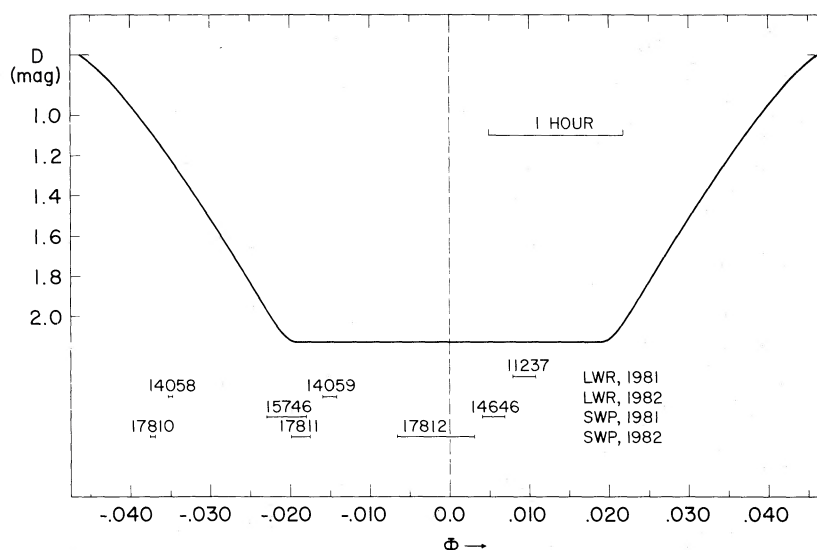


FIG. 7.—The distribution of the SWP and LWR exposures as function of phase, compared with a theoretical light curve during the primary eclipse.

against this hypothesis (Plavec, Weiland, and Koch 1982), the most powerful among them being that in SX Cassiopeiae, the emission lines are weakened during the eclipse of the primary (accreting) component. But the observed weakening was only by about 40%, and with the uncertainty in measuring the line fluxes, some doubts could have remained. The case of U Cephei is much more convincing: the emission lines do undergo eclipses and are considerably weakened during totality. Thus, *the line-emitting region surrounds the primary star*. We will now attempt to localize it more precisely.

We will use Figures 8 and 9. Figure 8 shows the projection of the disks of the two stars onto the plane of the sky, while Figure 9 presents the projection onto the orbital plane. These two projections are very nearly perpendicular to each other, since the orbital inclination is close to 90° . The pretotality spectrum, SWP 17811, is labeled 82 *P* in Table 5, while the totality spectrum (SWP 17812) is labeled 82 *T*. The former exposure covered phases 0.9803–0.9825, the latter 0.9934–0.0031. Because of the rather fast relative motion of the two stars, a certain amount of smearing of their relative

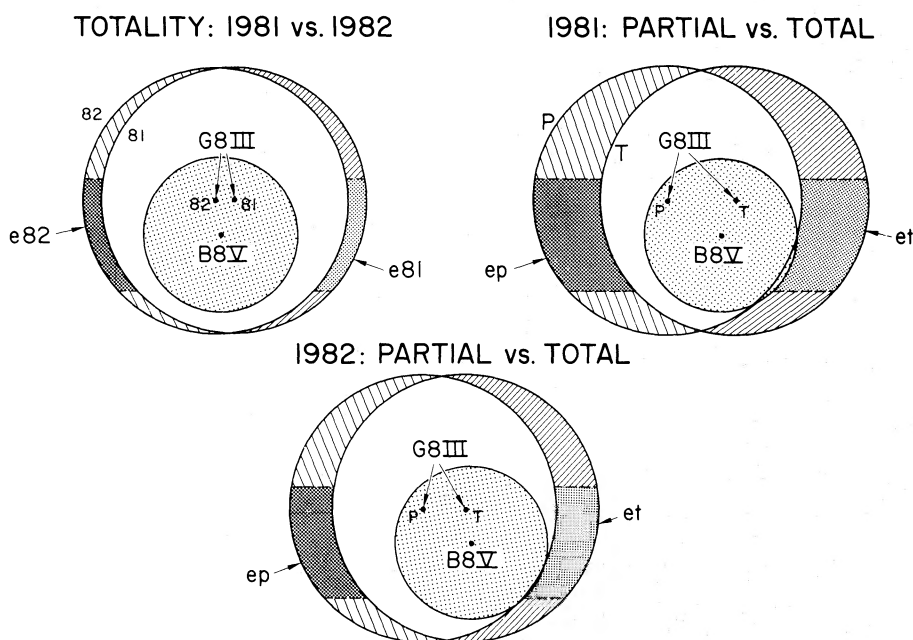


FIG. 8.—Projection of the disks, upon the plane of the sky, of the two stars during some of the IUE exposures. See text for explanations.

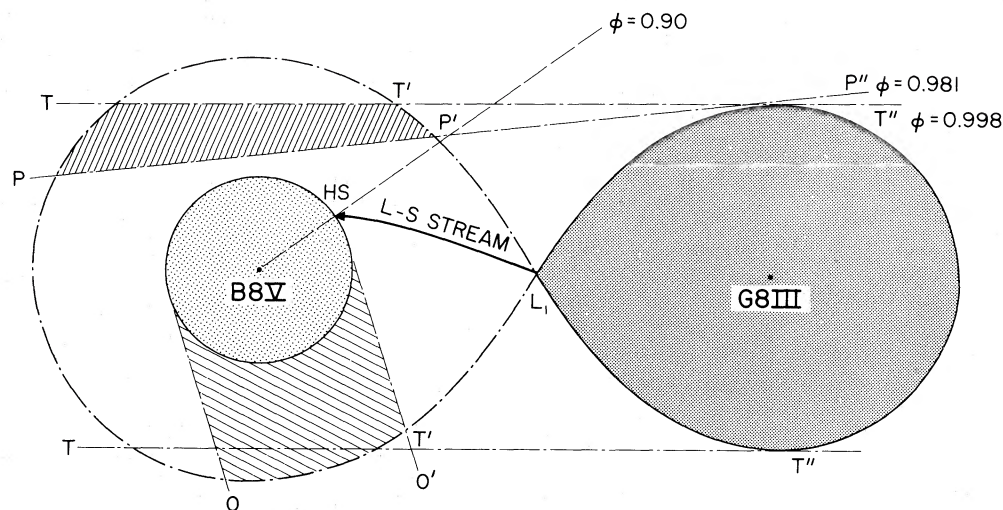


FIG. 9.—The system of U Cephei viewed pole-on. Boundaries of the observable regions are indicated for two eclipse spectra and for the high-dispersion spectrum SWP 14624 (0 0').

positions exists, which is hard to take into account. We will represent the two exposures by their mid-exposure phases, which are 0.91814 and 0.9981, respectively.

In the three diagrams in Figure 8, the position of the center of the B8 star is fixed, and the star's disk is dotted. The G8 star is supposed to move from left to right, and the two respective positions for each pair of spectra to be compared are marked. In the lower diagram, which applies to the case we are considering, we distinguish its *P* and *T* positions. Areas eclipsed for one spectrum but visible for the other spectrum are hatched. To the left is the area *ep* eclipsed during the *P* exposure, while to the right is the area *et* eclipsed during the *T* exposure. The region inside the hatched cusps was eclipsed in both cases, while the region outside was visible in both cases. Thus, the difference in line intensities between *P* and *T* implies that much more radiation was coming from *et* than from *ep*. Note that *et* is at the trailing hemisphere of the hotter star, where the incoming gas stream begins to bend around the B star.

We will attempt to define the most important radiating area more precisely. It is unlikely that the line-emitting region is spherically symmetrical. All circumstellar structures in interacting binaries tend to form disks concentrated toward the orbital plane. However, as I will argue in § VII, in U Cephei the structure cannot be geometrically thin perpendicular to the orbital plane. A reasonable estimate is that the "disk" has a semi-thickness of about 80% of the radius of the B star; it cannot be much thinner, as will be shown in § VII and as was also argued convincingly by Olson (1980a). Thus, we are left with the heavily hatched regions in Figure 8.

The heavily hatched region *et* radiated at least one-half of all the line intensity observed in spectrum *P*. It lies quite close to the surface of the B star, within $0.0\text{--}2.2 R_{\odot}$, or $0.0\text{--}0.8 R(B)$. The region so defined has a surface area

of about $3 \times 10^{22} \text{ cm}^2$. Figure 9 shows the configuration projected into the orbital plane. The two straight lines *TT'* and *PP'* delimit the region that projects into area *et* in Figure 8. A very crude estimate puts the depth of this region, along the line of sight, at some $5 R_{\odot}$; then the radiating volume is on the order of $1 \times 10^{34} \text{ cm}^3$.

As to the electron density in this region, we have no quite reliable clues. Figure 6 shows that the 1982 pre-totally spectrum is much richer in emission lines than the totality spectrum. Among the additional lines, two may be interpreted as being due to intercombination lines: The broad emission at $1747\text{--}1751 \text{ \AA}$ could be N III], and the peak at $1485 \text{ \AA}$ could be N II]. The former multiplet has a critical electron density of about $5 \times 10^{10} \text{ cm}^{-3}$; the latter line, $4 \times 10^{11} \text{ cm}^{-3}$. On the other hand, C III] $1909 \text{ \AA}$ is absent. With Si III] $1892 \text{ \AA}$ badly blended with Fe III $1895 \text{ \AA}$, such an interpretation of the pre-totally spectrum is compatible with an electron density of about $4 \times 10^{10} \text{ cm}^{-3}$. The corresponding emission measure would be $\text{EM} \approx 2 \times 10^{55} \text{ cm}^{-3}$. I think that this interpretation is improbable, though. It would probably require a circumstellar envelope or disk with radially increasing density, which is unlikely but perhaps not impossible if radiation pressure plays a dominant role near the stellar surface. But the 1982 pre-totally spectrum is rather similar to low-dispersion spectra of β Lyrae, where we also observe an increase of flux longward of about $1700 \text{ \AA}$. High-dispersion spectra of β Lyrae show that this increase is due to an accumulation of numerous emission lines, mainly those of Fe III and Ni III. For both U Cephei emissions, the one at $1485 \text{ \AA}$ and the complex at $1747\text{--}1751 \text{ \AA}$, there exist alternative explanations in terms of metallic lines. In that case, our previous estimate of $n_e \geq 5 \times 10^{11} \text{ cm}^{-3}$ holds.

Let us comment briefly on the other two pairs of spectra for which the corresponding relative star positions are visualized in the upper diagrams of

Figure 8. For the two totality spectra (*upper left*), the difference in position is small. Nevertheless, the 1981 spectrum shows definitely stronger emissions, on the average by about 36%. This is probably a non-periodic effect. However, if interpreted as an eclipse effect, it says that the region *e82* is emitting somewhat more strongly than *e81*. Both regions lie roughly half a radius above the surface of the B8 star, but the more strongly emitting region would this time be at the approaching hemisphere. A comparison of the 1981 pretotality spectrum with the 1981 totality spectrum (*upper right diagram*) essentially supports our first finding, namely, that the trailing (right) side of the disk emits more. The difference in line intensities between the two spectra is smaller, though—only 26% on the average, with a large scatter—and the radiating regions defined by the two positions of the eclipsing star are larger, so that this comparison does not add much to our knowledge.

VII. THE ABSORPTION SPECTRUM OF U CEPHEI

Outside the eclipse, it is possible to observe the far-ultraviolet spectrum of U Cep in the high-dispersion mode of the *IUE* spectrograph. Such observations were already made and described by Kondo, McCluskey, and Harvel (1981), who obtained three SWP spectra at phases 0.184, 0.680, and 0.903. Our high-dispersion SWP spectrum was obtained at phase 0.199. The essential features in the spectrum were described by Kondo *et al.* as follows. Most lines are shallow and broad. This is interpreted, in accord with the optical spectrum, to imply a high velocity of axial rotation, above 300 km s⁻¹. Several sharp components are obviously of interstellar origin; Kondo *et al.* list four of them, but we see at least seven and list them in Table 6. The mean of the radial velocities derived for these interstellar components, which were obtained in a rather crude way, is -21 km s⁻¹ in our spectrum, while Kondo *et al.* obtained -9 km s⁻¹.

The most remarkable feature pointed out by Kondo *et al.* is the presence of the absorption lines such as Si IV (1) and C IV (1). At the photospheric temperature appropriate for the spectral type of U Cep A, C is primarily C⁺, and Si is primarily Si⁺⁺. Therefore, the Si IV and C IV lines must originate in a different and much hotter region (comparable to a B0-B1 star);

Kondo *et al.* suggest the existence of a “pseudophotosphere” along the equatorial region of the primary star. They found an equivalent width on the order of 0.7–0.9 Å for Si IV (1) and C IV (1), about 1–2 Å for the Si II (4) lines near 1260 Å, and between 2 and 3 Å for C II (1). The ranges given reflect not only the uncertainty of measurements in the heavily blended spectrum, but also genuine variations from phase to phase. Our measurements roughly agree with their values (cf. Table 3).

But we want to go one step further in the interpretation of the absorption spectrum. We claim that all the lines listed in Table 3 as emission lines have their counterparts in the absorption spectrum, and, most importantly, that the corresponding lines in the absorption spectrum have profiles distinctly different from the typical photospheric lines. They do not possess the bell-shaped broad and shallow profiles typical for rapid rotation. As an example, see Figure 10, where the Fe III (34) line at 1895 Å is one of these “hot” lines which are not formed in the photosphere of the B8 star. No doubt most of the lines, like those of Si II or C II, are formed partly in the real photosphere and partly in the “hot” region, since Kondo *et al.* found some relation to the orbital motion of the B8 star. In fact, they claim that the C II (1) line shows an orbital velocity almost the same as that of the B8 star, which is surprising, since it is fairly strong as an emission line. Even for those mixed lines, though, the profiles are sufficiently different from the broad shallow photospheric profiles. Thus, we have here an opportunity to check on our line identifications in Table 3, since the absorption spectrum has a much better resolution. This was fully used in preparing Table 3, where in the last columns the equivalent widths (EW) of the lines are given, as well as their central residual intensities (*r_c*).

But, most importantly, in the high-dispersion absorption spectrum we are getting another look at the emission-line region: this time we are looking through it toward the star. It is unlikely that we are observing exactly the same region in space, since the “hot” layer is probably not azimuthally symmetrical with respect to the star, so we see different segments of it. Nevertheless, we may attempt to draw some important quantitative conclusions about this layer. For this purpose, we will consider only those lines in which the contribution from

TABLE 6
INTERSTELLAR ABSORPTION LINES SEEN IN U CEPHEI

$\lambda_{\text{obs}}(\text{\AA})$	$\lambda_{\text{lab}}(\text{\AA})$	ID	V_r (km s ⁻¹)	Comments
1253.7	1253.812	Si II (1)	-27	
1260.3	1260.421	Si II (4)	-29	Fe II (9) may contribute
1302.2	1302.169	O I (2)	+7	Unlikely to be Si III (4) as Kondo <i>et al.</i> 1981 suggest
1304.2	1304.372	Si II (3)	-40	
1334.55	1334.532	C II (1)	+4	
1335.5	1335.703	C II (1)	-46	
1526.65	1526.708	Si II (2)	-11	

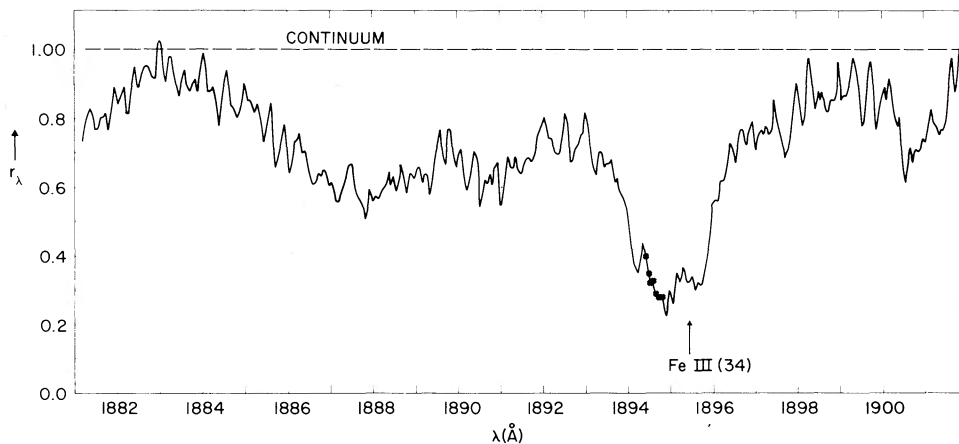


FIG. 10.—The line Fe III (34) 1895 Å, originating in the “hot” region, has a profile different from typical photospheric lines, represented here mainly by the Fe II (125) line at 1888.7 Å.

the genuine photosphere is nil or at least small: namely, the lines of N v (1), C iv (1), and Si iv (1). Assume that the absorptions are formed in a scattering “reversing” layer above the photosphere. Then the central residual intensities give us the optical thickness of the reversing layer, τ_0 (Ambartsumyan 1958):

$$\tau_0 = \left(\frac{4}{3}\right)(1/r_c - 1). \quad (3)$$

Assuming a uniform reversing layer, we can write

$$\tau_0 = \alpha H R_i A_{ei} N_H = 0.8 \alpha H R_i A_{ei} N_e, \quad (4)$$

where α is the absorption coefficient per atom, H is the geometrical thickness of the reversing layer, A_{ei} is the fractional abundance of the element with respect to hydrogen, and R_i is the fraction of the atoms of that element which are in the ionization state i . Since we are dealing with resonance lines, we assume that all the $(N_H A_{ei} R_i)$ absorbing atoms are in the ground level. For the absorption coefficient at the line center we will adopt

$$\alpha = [(\pi)^{1/2} e^2 / mc^2] \lambda^2 f / \Delta \lambda_D, \quad (5)$$

with

$$\Delta \lambda_D = (\lambda/c)(2kT_e/\mu m_H + v_t^2)^{1/2}. \quad (6)$$

It is justified to assume considerable turbulent velocity v_t , since the lines are very broad, yet they do not have the bell-shaped profiles seen in pure photospheric lines of U Cephei. The profiles are so broad that turbulence must be the dominant factor in formula (6), and the Doppler width is rather insensitive to the electron temperature of the gas. The gas temperature enters (as a much more decisive factor) into the ionization fractions R_i . Assuming solar abundances, we have calculated the ionization fractions R_i under the assumption of collisional equilibrium, using the recent results of Shull and Van Steenberg (1982).

The observed line profiles must be corrected for the fact that the circumstellar disk is unlikely to cover the

whole visible disk of the B star. In fact, the “canonical” picture of an accretion disk postulates a relatively thin disk located in the orbital plane. It is easy to show that in U Cep (and in other W Serpentis systems as well) the accretion disk must be very much thicker and its thickness comparable to the radius of the star. Assume that only a fraction B of the visible disk is obscured by the absorbing layer. Then the central residual intensity in the absorption lines will be given, instead of by equation (3), by the following relation:

$$r_c = (1 - B) + B/(1 + \tau_0). \quad (7)$$

Even for completely dark line centers ($\tau \rightarrow \infty$) within the absorbing layer, the observed central residual intensities cannot be smaller than $(1 - B)$. Therefore, the observed $r_c = 0.28$ for the Si iv 1394 Å line implies that $B \geq 0.72$. This line, being unusually deep, may not be quite reliable. But the C iv lines still demand $B \geq 0.61$. Since very large optical depths in the reversing layer are unlikely, we must conclude that its geometrical thickness must be comparable to the star’s radius. Olson (1980a) has argued that the transient disks occasionally observed, which are optically thick in the continuum, have perpendicular thickness comparable to the radius of the B star.

Out of the six absorption lines available, only four could be used for line profile fitting; N v 1238 Å and C iv 1551 Å are too distorted by blends. In the model fitting, three parameters have been varied over a wide range of values: the central optical depth τ_0 and the turbulent velocity v_t characterizing the intrinsic profile and the rotational velocity v_{rot} that broadens the intrinsic profile. It is known that the effects of turbulence and a moderate rotational velocity are not easy to separate, and both in turn strongly affect τ_0 in the fit. Combined with the rather low signal-to-noise ratio, uncertainties in the instrumental profile and calibration, and danger of blends, all these factors make the results given in Table 7 rather crude. But the fairly large optical thickness in the lines has been established without doubt, and the broad but distinctly different line profiles make the assumption

TABLE 7
IMPORTANT PARAMETERS OF SELECTED ABSORPTION AND EMISSION LINES

LINE	ABSORPTION LINES			EMISSION LINES	
	τ_0	v_t^a	v_{rot}^a	P_{obs}^b	ϵ^c
N v 1240 Å	3.8	100	230	6.1×10^{30}	2.935×10^{-25}
C II 1336 Å	...	...	...	8.3×10^{30}	5.953×10^{-25}
Si IV 1394 Å	35.0	100	170	2.7×10^{31}	3.876×10^{-25}
Si IV 1403 Å	33.0	90	250		
C IV 1548 Å	30.0	80	240	3.2×10^{32}	1.417×10^{-23}

^a Units: km s^{-1} .

^b Total power. Units: $\text{ergs cm}^{-2} \text{s}^{-1}$.

^c Emissivity at $T_e = 10^5 \text{ K}$ and $n_e = 1 \text{ cm}^{-3}$, with $Y = 0.75$.

of a large turbulent velocity also safe. Thus, qualitatively, our results are well established. Crawford (1981) concluded, from his analysis of the H α emission line in U Cep, that the large width of the line must be due to supersonic turbulence amounting to about 100 km s^{-1} . This agrees surprisingly well with our results.

Table 7 lists, for each of the four lines studied, the central optical thickness of the intrinsic profile τ_0 , the turbulent velocity v_t , and the rotational velocity v_{rot} (the last two in km s^{-1}). It is not clear if all the four lines are formed in the same region, but the good agreement in the turbulent velocity makes it possible. The scatter in the rotational velocity probably indicates the uncertainty with which the observed wings of the lines are defined. There is little doubt that the rotational velocity is substantially lower than the value usually given for the photosphere of the star; the meaning of this result is not clear. Table 8 gives the values of the effective electron column density (HN_e) as defined by equation (4), for various values of electron temperature T_e . Both the column densities (in cm^{-2}) and electron temperatures (at the top of the table) are expressed in logarithms.

Let us attempt to evaluate the results under the assumption that all the lines are formed in the same region of constant temperature T_e . Table 8 shows that it is impossible to find a temperature for which the four lines would give at least approximately concordant data. The main source of the discrepancy is no doubt the oversimplified assumption of a homogeneous isothermal reversing layer. Interestingly, the Si IV effective thicknesses lie between those for C and N, and the disagreement can be reduced if we assume that the circumstellar

material surrounding the hotter component of U Cephei has been (at least partially) processed by the CNO cycle. We assumed solar composition, which according to Lang (1978) implies the abundance ratios C:N:O = 3.7:1:6.5. If the CNO cycle has been operating for a time sufficiently long to establish the C/N equilibrium but not the N/O equilibrium, the abundance of N can be raised by a factor of up to 4.6, while the abundance of carbon can be reduced by a factor of up to 39.4 (Clayton 1968). In this way, the effective geometrical depths of the reversing layer can be brought to a somewhat better agreement, in particular for electron temperatures near $\log T_e = 5.0$. The mean column electron density is then of the order of $5 \times 10^{21} \text{ cm}^{-2}$.

As mentioned in § VI, the absence of the intercombination lines indicates an electron density of at least $5 \times 10^{11} \text{ cm}^{-3}$. If so, then the geometrical thickness of a homogeneous absorbing layer would be only $10^{10} \text{ cm} = 0.14 R_\odot$. But the observed extent of the circumstellar radiating region is much larger. The region marked *et* in Figure 8 has a radial thickness of $2.2 R_\odot = 1.5 \times 10^{11} \text{ cm}$. This would in turn require an electron density as low as 10^{10} cm^{-3} . This is a serious discrepancy. Perhaps the highly ionized material exists in blobs, or in thin sheets or filaments formed as a consequence of numerous shocks. If so, then the filling factor would be small, and the discrepancy could be removed.

If the electron density is indeed as high as 10^{12} cm^{-3} or higher, the ionization equilibrium calculations may require modifications. The available observational data do not yet warrant any more sophisticated treatment.

TABLE 8
THICKNESS ESTIMATES OF THE "HOT" CIRCUMSTELLAR LAYER IN U CEPHEI
(NO CNO PROCESSING ASSUMED)

LINE	LOG (HN_e) (cm^{-2})					
	LOG $T_e = 4.7$	4.8	4.9	5.0	5.1	5.2
N v 1240 Å	29.20	26.36	23.86	22.21	20.99	20.21
Si IV 1394 Å	23.83	23.03	22.38	21.93	21.93	22.26
Si IV 1403 Å	24.08	23.28	22.63	22.18	22.18	22.50
C IV 1548 Å	22.28	21.08	20.20	19.79	20.13	20.71

VIII. NATURE OF THE LINE-EMITTING REGION

The amount of line radiation reaching us from the region whose visible face has been labeled *et* in Figure 8 can be estimated from Table 5. The excess flux seen during the 1982 partial eclipse (82 *P*) over that seen at the total eclipse (82 *T*) is the difference of the fluxes coming to us from the regions *et* and *ep*. The radiating volumes are roughly equal, and we can estimate that the emitted powers scale approximately as $P(ep) \approx 0.4P(et)$. The resulting estimates for several crucial lines are entered in Table 7.

We will now attempt to predict emission-line intensities coming from the region *et*, assuming at first that the emission lines are optically thin. We assume that the ionization equilibrium is collisional, and we use the recent ionization tables of Shull and Van Steenberg (1982) to obtain the ionized fraction R_i for the ions of C, N, and Si. The emission-line intensities are calculated by means of the formula

$$P = \epsilon(T_e) N_e^2 V, \quad (8)$$

where V is the emitting volume, and ϵ is the emissivity per unit volume and electron density of 1 cm^{-3} , given by

$$\epsilon = 2.179 \times 10^{-15} g_{\text{eff}} f A_{ei} R_i \exp(-hc/k\lambda T)/(T_e)^{1/2}. \quad (9)$$

Owing to the limited and inaccurate data, we can consider here only the simplest case, namely, that of an emitting region homogeneous in electron temperature and electron density. In such a case, the ratio of the observed fluxes is the same as the ratio of the respective emissivities.

This ratio should be a function of only the electron temperature if we consider the two (blended) multiplets of carbon and calculate the intensity ratio of $P(\text{C IV})/P(\text{C II}) = P(\lambda 1549)/P(\lambda 1336)$. The observed ratio is 38. Computations show that this ratio is a very steep function of T_e and, in the vicinity of $T_e = 10^5 \text{ K}$, increases by an order of magnitude per $\Delta \log T_e = 0.1$. For $\log T_e = 5.0$ the ratio is 24, and for $\log T_e = 5.1$ it is already 192. If our assumption of a homogeneous emitting region is correct, then the electron temperature must be very close to $1 \times 10^5 \text{ K}$, in perfect agreement with our findings from the absorption lines.

The observed ratio $P(\text{C IV})/P(\text{N V}) = P(\lambda 1549)/P(\lambda 1240)$ is 51.5. This ratio depends on both the electron temperature and the relative abundances of carbon and nitrogen. We will express the latter dependence by means of a parameter Y , which is the fraction of the emitting material that has been processed by the CNO cycle. We can write, using the data from Clayton (1968),

$$A_N = (A_N)_0 (1 + 3.6085Y), \quad (10)$$

$$A_C = (A_C)_0 (1 - 0.9746Y), \quad (11)$$

where the zero subscripts indicate the unprocessed abundances taken from Lang (1978). Again, since the

ionization fraction of N^{+4} increases steeply with temperature in the vicinity of $T_e = 10^5 \text{ K}$, the electron temperature is determined rather sharply and should lie within the narrow range $4.9 \leq \log T_e \leq 5.1$. A full agreement with the observed ratio is reached at $T_e = 1 \times 10^5 \text{ K}$ if we assume that $Y = 0.75$; i.e., that 75% of the circumstellar material has been processed in the CNO cycle. The observed ratio of intensities almost certainly does not represent the optically thin case, which would probably give an even larger ratio. No larger deviation from the assumed electron temperature would result. But the CN-processed fraction Y is very sensitive to the adopted temperature, and one can get the observed intensity ratio with no CN processing at an only slightly higher T_e , about $1.2 \times 10^5 \text{ K}$. Thus the emission-line intensities do not seem to provide any additional strong argument in favor of the CNO processing of the circumstellar material.

If we now accept $T_e = 1 \times 10^5 \text{ K}$ and $Y = 0.75$ as the best available estimates, then the emissivities given in Table 7, combined with the observed line intensities, give us an estimate of the emission measure $\text{EM} = N_e^2 V$. We obtain an average value of $\text{EM} \approx 3 \times 10^{55} \text{ cm}^{-3}$. Using the lower limit on the electron density obtained from the absence of the intercombination lines, $n_e = 5 \times 10^{11} \text{ cm}^{-3}$, we find that the emitting volume is $V = 1.2 \times 10^{32} \text{ cm}^3$. This value is 80 times smaller than our estimate from the geometry of the system, which was $V = 10^{34} \text{ cm}^3$. If we, in turn, adopt this value, then a significantly lower electron density results, namely, $n_e \approx 6 \times 10^{10} \text{ cm}^{-3}$. We have encountered the same dilemma as in the case of the absorption lines. The emission intensity is no doubt reduced significantly, compared with the optically thin case, by the optical thickness of the emitting region. From the geometrical arguments, we can estimate that the depth of the volume *et* is about twice its radial thickness, and the same factor may enter the optical depths. But apart from this reduction in the emission power, it again appears likely that the volume actually participating in the emission may be significantly smaller than the available volume: the idea of blobs or filaments comes to mind again. Naturally, the whole concept of a homogeneous emitting region may be considerably in error. Perhaps the active region is structured similarly to a chromosphere, with a steep temperature gradient. Only refined observations can decide. At the present, the evidence is rather contrary to a large thermal gradient. As the partial eclipses of the emitting region indicate, it is rather extended around the accreting star. Any strong thermal stratification should show in strongly variable intensities of the emission lines with phase, and no such effect has been observed. Moreover, if each line were formed, as in the solar transition zone case, almost exclusively in a temperature range in which the corresponding ion is most abundant (that is, the N V lines would be formed at $T_e \approx 2.2 \times 10^5 \text{ K}$), the effective thicknesses of the line-emitting layers would be even much smaller, of the order of 10^7 cm . This is acceptable in the atmosphere of a dwarf star but not in our case.

The absence of the He II 1640 Å line in emission (and most likely also in absorption) poses an interesting problem. The power emitted by the line can be written, in the optically thin case, as

$$P(\text{He}) = \epsilon(\text{He}) N_e^2 V, \quad (12)$$

with

$$\epsilon(\text{He}) = 0.8 h \nu \alpha A_{\text{He}} R_2, \quad (13)$$

since this line—an analog of H α —results from recombination of He⁺. Here $\alpha = 68 \times 10^{-14} (T/10^4)^{-0.88}$ is the recombination coefficient, according to Osterbrock (1963), Spitzer (1978), and Jameson, King, and Sherrington (1980). For $T_e = 10^5$ K, we get $\epsilon(\text{He}) = 4.37 \times 10^{-26}$ cm³. Compared with the lines we discussed above, the intensity of the He II 1640 Å line comes out to be, in the optically thin case, 11% of that of Si IV, or 7% of C II 1336 Å. This is within the limits of observational uncertainty. Thus it seems that at least in the case of U Cephei we can explain the absence of the He II lines without additional assumptions.

IX. DISCUSSION

When the emission lines in the interacting binaries of the W Serpentis type were first discovered in 1978, several colleagues suggested that the lines are formed in the chromosphere (actually in the transition region) of the late-type giant which is the secondary component of many interacting binaries such as U Cep. But there is no late-type giant in β Lyrae, V356 Sagittarii, or V367 Cygni. Arguments against such a model being applicable to the W Serpentis stars were presented in our article on SX Cas (Plavec, Weiland, and Koch 1982). This debate seems decided by the very clearly observed partial eclipses of the line-emitting region in U Cephei. As an additional argument, consider the surface fluxes obtained from the data in Table 3 by dividing the total fluxes by the surface area of either component. Such surface fluxes for the strongest lines are on the order of 10^6 – 10^7 ergs cm⁻² s⁻¹, i.e., two to three orders of magnitude larger than in Capella or in the most active known RS CVn stars (cf. Linsky *et al.* 1978). Only one category of stars yields very similar surface fluxes, namely, the T Tauri stars (Imhoff and Giampapa 1982). But T Tauri stars possess no quiescent chromospheres. Rather, there exists a vigorous interaction in their circumstellar envelopes between the infalling material, the accreting star, and material that is being ejected from the star's vicinity. It should not be surprising that the spectral phenomena in T Tauri stars and in W Serpentis stars bear certain resemblances: after all, fundamental underlying processes may be rather similar. In T Tauri stars, the material is being accreted from a circumstellar cloud at a rate of perhaps $10^{-7} M_\odot$ yr⁻¹, while in the Serpentids the source is the secondary component, and the influx rates may be several orders of magnitude higher.

We have established that the line-emitting region

surrounds the hotter star. It is almost certain that the energy powering and ionizing this region comes ultimately from the gravitational energy released in the process that is generally called accretion. How is it released?

The theory of gas flow in semidetached close binaries, as developed by Lubow and Shu (1975), predicts that the stream leaving the G8 star should impact on the B8 component with a velocity of ~ 620 km s⁻¹ at an angle of $\sim 40^\circ$ to the normal. The trajectory is shown in Figure 9 as the "L-S stream." The radial component of the impact velocity is about 480 km s⁻¹; the tangential component about 390 km s⁻¹. If the radial component is completely dissipated, then 1 gram of the impacting material delivers 10^{15} ergs of energy. The rate of mass transfer is very uncertain, but it probably lies between 10^{-5} and $10^{-6} M_\odot$ per year. Then the rate of energy deposition lies between 6×10^{34} and 6×10^{35} ergs s⁻¹, or between 15 and 150 $L_\odot$. Only a tiny fraction of this energy is converted into flux in the ultraviolet emission lines, as has been established in § V: probably no more than 1%. Thus it may still hold true that most gravitational energy is released at the hot spot formed when the stream penetrates the photosphere of the B8 star. But part of the material and of the energy is quickly redistributed along the whole circumference of the accreting star and forms the "hot" layer which, in U Cephei, has now been observed both in emission and in absorption. It is probable that the rapid rotation of the B8 star (which in turn was most likely generated by the oblique infall of the stream) contributes to rapid redistribution, since the period of axial rotation is only 10.6 hr. It has been suggested (Wilson and Twigg 1980) that the rapidly rotating components in semidetached binaries are actually rotationally unstable and hence are unable to accommodate the in-flowing material. Rapid rotation of the accreting star in U Cep may play an important role. However, Peters and Polidan (1983) find similar highly ionized circumstellar layers in other Algol-type eclipsing binaries, such as U CrB, AU Mon, and CX Dra, and Kaitchuck and Honeycutt (1982) report the same for RW Tau; and in these, the accreting star is not rotating very rapidly. We do not understand what mechanism creates the circumstellar "hot" turbulent layer, but there is little doubt that such a layer exists in many interacting semidetached binaries.

The following properties have been established for the "hot" layer in U Cephei by our observations:

1. The layer is optically thick in the absorption lines of N V (1), Si IV (1), and C IV (1), when viewed in the radial direction. This almost certainly means that the emission lines are also optically thick.
2. The layer probably surrounds the whole star along the equator, since the same absorptions have been observed at various phases, although with variable intensity.
3. The thickness of the "hot" layer perpendicular to the orbital plane cannot be much smaller than the radius of the accreting B star: we are not dealing here with a thin accretion disk.

4. The radial extent of the layer is very uncertain. It may actually be clumpy, perhaps formed by thin shocked regions.

5. It is probable—but far from proven yet—that the circumstellar material in U Cep does not have the solar composition, but rather that it has been partly processed by the CNO cycle, which shifted the equilibrium between the C and N abundances in favor of nitrogen.

A certain similarity to the T Tauri stars, mentioned above, makes it worthwhile to consider the model recently advanced by Ulrich (Ulrich *et al.* 1983): Because of the fluctuations in the influx rate, no stable shock is formed where the material should impact on the star, but rather a highly turbulent region (the *turbosphere*) develops, in which the gravitational potential energy is released in a series of small shocks. Ulrich proposes that such a turbosphere would be characterized by equipartition of energy between thermal and turbulent, so that the relation between the turbulent velocity v_t and the electron temperature of the gas T_e reads

$$kT_e = \mu m_H v_t^2; \quad (14)$$

i.e., the turbulent velocity (in km s^{-1}) will be given by

$$v_t = 0.113(T_e)^{1/2}. \quad (15)$$

The energy released in the turbosphere will eventually be converted into optically thin bremsstrahlung of hydrogen, i.e., spread over a much wider range of wavelengths than is the case of blackbody radiation.

This model is attractive and may be in principle

applicable to the W Serpentis stars, but the numerical results we obtained for U Cep do not agree with the turbospheric model. For an electron temperature close to 10^5 K, the postulated equipartition of energy leads, according to equation (15), to a turbulent velocity of only 40 km s^{-1} , while the evidence accumulated in this article, as well as by Crawford (1981) and Peters and Polidan (1983), requires a value close to 100 km s^{-1} .

Although it now appears fairly safely established that the line-emitting region surrounds the accreting star, it is appropriate to conclude with a word of caution. It has recently been realized that the outer layers of single stars are very complex, and that stellar wind is often present. It is very likely that the corresponding structures and phenomena in interacting binary stars will be far more complicated. A wind from the component that fills the critical Roche lobe may play a more important role than we are able to recognize at this stage.

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