STELLAR WINDS AND MASS-LOSS RATES FROM Be STARS

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Resonance-line profiles of SiIII and SiIV lines in 22 B and Be stars have been analyzed in the derivation of mass-loss rates. Of the 19 known Be or shell stars in the sample group, all but one show evidence of winds. It is argued that for stars of spectral type B1.5 and later, SiIII and SiIV are the dominant stages of ionization, and this conclusion, together with theoretical fits to the line profiles, leads to mass-loss rates between 10^{-11} and 3×10^{-9} for the stars. The rate of mass loss does not correlate simply with stellar parameters, and probably is variable with time. The narrow FeII shell lines often seen in the ultraviolet spectra of Be stars may arise at low levels in the wind, below the strong acceleration zone. The mass-loss rates from Be stars are apparently insufficient to affect stellar evolution.

1. INTRODUCTION AND PREVIOUS RESULTS

The Be stars show a wide variety of phenomena, many of them seemingly unrelated or even in conflict with each other. No two stars behave exactly the same, and many, if not all, are variable with time.

In the 1970's a new ingredient was added to the Be star stew: stellar winds were discovered, first by Bohlin (1970), who found P Cygni profiles in the ultraviolet spectrum of γ Cas, and later by Snow and Marlborough (1976) and Marlborough and Snow (1976), who found asymmetric ultraviolet profiles indicative of high-velocity winds in a number of these objects. Later, Lamers and Snow (1978) added to the list of Be stars known to have winds, and it became apparent that high-velocity outflow from the stars has to be considered along with all the other phenomena associated with Be and shell characteristics.

Until now, very few attempts had been made to quantitatively determine the rates of mass loss associated with the winds from Be stars. Snow and Marlborough (1976) crudely estimated that 59 Cyg is losing matter at a rate between 10^{-10} and 10^{-9} $M_\odot$ yr^{-1}, but the first real quantitative analyses were carried out by Bruhweiler, Morgan, and van

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der Hucht (1978), who used metastable FeIII lines in the near-ultraviolet spectrum of ϕ Per to estimate $\dot{M} = 5 \times 10^{-11}$ for that star; and by Hammerschlage-Hensberge et al. (1980), who analyzed ultraviolet resonance lines in the spectra of γ Cas and X Per to derive values of $\dot{M}$ equal to 7×10^{-9} $M_\odot$ yr^{-1} and 1×10^{-8} $M_\odot$ yr^{-1}, respectively.

In the present study, Copernicus ultraviolet spectra of 22 early-to-mid-B objects, 19 of which have been recognized as Be or shell stars, have been studied with the aim of deducing mass-loss rates and then determining, if possible, how these rates are related to other stellar and circumstellar properties of the stars.

2. THE OBSERVATIONAL DATA

The Copernicus data used in this analysis were obtained primarily at a resolution of 0.2 A, although in a few cases a higher resolution of 0.05 A was used. All photometric corrections for stray light and backgrounds were carried out using standard procedures, and together are thought to introduce errors in intensity of no more than 20% of the continuum level. The details of the data reduction are outlined in a separate paper (Snow, 1981).

The stars observed were not selected in any systematic way, but simply represent the sampling of Be stars that happened to be observed by Copernicus at the appropriate wavelengths. In some cases the data were originally obtained specifically for studies of either stellar winds or Be stars, and nearly all the stars have been discussed in other publications by this author (Snow and Marlborough, 1976; Marlborough and Snow, 1976; Lamers and Snow, 1978; Snow, Peters, and Mathieu, 1979) or by others (e.g. Marlborough, 1977; Peters, 1976, 1979; Heap, 1976; Burton and Evans, 1976).

Table 1 lists the stars, along with assorted basic data taken from the references just mentioned. The values of v sin i are from Slettebak et al. (1975), or were scaled to their system, using the standard correlation. The T_{eff} entries are based on the listed values of R_* and M_{bol}, which in turn are derived from standard values given in Code et al. (1976) for R_*; and from the compilations of Lesh (1968,1972), along with bolometric corrections from Code et al. for M_{bol}.

A number of the stars in Table 1 are known to undergo time variability. Particularly well known to have variable winds are γ Cas (Henrichs, 1980), 59 Cyg (Doazan, Kuhi, and Thomas, 1980; Marlborough and Snow, 1980; and Doazan et al., this volume), and δ Cen (Snow, Oegerle, and Polidan, 1980). This variability, which is probably present in many of the other stars as well, will be mentioned again.

TABLE 1. THE STARS AND THEIR PROPERTIES

Star	HD	Spect.	V	E(B-V)	M_{bol}*	R_*⁺ ($R_\odot$)	T_{eff} (°K)	v sin i** (km s^{-1})	$\dot{M}$ ($M_\odot$ yr^{-1})
γ Cas	5394	B0.5 IVe	2.58	0.08	-6.85	8.5	28,500	230:	6.0×10^{-11}
δ Sco	143275	B0.5 IV	2.33	0.16	-7.25	8.5	31,300	150	3.0×10^{-11}
η Ori	35411	B0.5 Vnn	3.42	0.11	-6.5:	8.5	26,300	40	9.3×10^{-11}
25 Ori	35439	B1 Vn	4.94	0.05	-5.71	8.0	22,600	260	2.2×10^{-9}
π Aqr	212571	B1 Ve	4.68	0.29	-6.08	8.0	24,600	300:	2.6×10^{-9}
	28497	B1.5 Ve	5.60	0.02	-5.21	7.5	20,800	290	7.8×10^{-10}
η Cen	127972	B1.5 V	2.3:	0.05	-5.21	7.5	20,800	260:	2.9×10^{-10}
59 Cyg	200120	B1.5 Ve	4.79	0.18	-5.21	7.5	20,800	350:	1.6×10^{-10}
δ Cen	105435	B2 IVne	2.59	0.12	-5.32	7.0	22,100	220:	7.4×10^{-11}
δ Cru	106490	B2 IV	2.80	0.00	-5.32	7.0	22,100	140	4.1×10^{-11}
δ Lup	138690	B2 IV	2.78	0.03	-5.32	7.0	22,100	210:	9.7×10^{-11}
ω CMa	56139	B2 IV-Ve	3.90	0.08	-4.76	7.0	19,430	80	1.4×10^{-10}
φ Per	10516	B2 IVep	4.06	0.22	-4.52	7.0	18,400	400	2.7×10^{-10}
μ Cen	120324	B2 IV-Ve	3.20	0.10	-4.72:	7.0	19,250	155	$<8.5 \times 10^{-11}$
υ Cyg	202904	B2 Ve	4.28	0.16	-4.52	7.0	18,400	200	6.3×10^{-11}
ζ Cen	121263	B2.5 IV	2.54	-.02	-4.22	6.8	17,400	160	3.8×10^{-11}
α Eri	10144	B3 Vp	0.48	0.04	-3.10	5.9	14,440	225:	1.3×10^{-10}
48 Per	25940	B3 Ve	4.03	0.17	-3.10	6.0	14,320	200	1.7×10^{-10}
ζ Tau	37202	B4 IIIp	2.95	-.01	-5.9:	6.0	27,300	300	2.4×10^{-10}
48 Lib	142983	B5 IIIp	4.87	0.06	-3.21	5.2	15,780	400:	6.7×10^{-11}
ψ Per	22192	B5 Ve	4.25	0.12	-3.61	5.2	17,300	350	7.6×10^{-11}
o And	217675	B6p	3.62	0.05	-1.6:	4.7	11,400	280	1.2×10^{-10}

* Values of M_{bol} were taken from previous studies of these stars, primarily Lamers and Snow (1978) and Snow, Peters, and Mathieu (1979).

\+ Values of R_* were based on standard calibrations of R_* with spectral type, most notably that of Code *et al*. (1976). In a few cases, stars from this list were specifically included in Code *et al*.

** Values of v sin i are from Slettebak *et al*. (1975), except for those that are underlined, which were taken from earlier compilations but converted to the system of Slettebak *et al*. using the correlation presented in their paper.

3. THE MASS-LOSS RATE ANALYSIS

Theoretical wind profiles from the atlas of Castor and Lamers (1979) were fitted to smoothed observational profiles of the SiIII and SiIV resonance lines at 1206 and 1393 A, respectively. In determining the fits, it was necessary to take into account the photospheric contributions to these lines, again using procedures given by Castor and Lamers.

The result of the fitting procedure was the derivation of the wind velocity law (which was not well-determined, but which also has very little effect on the mass-loss rate), a parameter that characterizes the changing ionization with height, and a total wind optical depth. These could then be combined with equations of radiative transfer and mass conservation to yield values of $\dot{M}$, the mass-loss rate.

A number of important uncertainties have to be kept in mind. Most significant of these is the ionization balance, since the mass-loss rates were based on the assumption that all the silicon is in the forms of SiIII and SiIV. It is well known that for the O and the earliest B stars, SiV is the dominant stage of ionization (e.g., Lamers, Gathier, and Snow, 1980), but the present data indicate that, for stars of type B1.5 and later, SiV is not important. Thus, the mass-loss rates for the program stars hotter than B1.5 may be underestimated due to the neglect of SiV, but for the cooler stars, this is probably not a serious omission.

When only SiIII or SiIV was observed, the mean SiIII/SiIV ratio for the winds of the other stars had to be assumed, introducing further potential errors.

Another major uncertainty was the measured value of the terminal velocity, which affects both the derived wind parameters from the profile fitting, and the calculation of $\dot{M}$. Detailed analysis shows, however, that the two effects compensate each other, and that the errors introduced into $\dot{M}$ by uncertainties in the terminal velocity are of the order of 30% or less, even if the velocity is underestimated by a factor of more than 2.

Other uncertainties arise from possible errors in the value of R_* that are adopted (a relatively minor effect) and from the possibility that the winds are not spherically symmetric. Furenlid and Young (1980) argue on the basis of Hα profiles that the winds are confined to the equatorial plane. Since the low v sin i stars in this sample show evidence for winds in the ultraviolet, this conclusion cannot be strictly true, but substantial enhancement in the equatorial plane is still possible. An attempt was made, following the method outlined by Hutchings (1976) and by Sonneborn and Collins (1977) to determine the stellar inclination angles from the visible and ultraviolet line widths, so that possible latitude-dependence of the winds could be assessed, but the data proved inadequate for this. Therefore, no sensible manner of allowing for latitude-dependence was found, and the rates of mass loss were derived on the assumption of spherical symmetry.

4. THE RESULTS

The derived values of $\dot{M}$ are listed in Table 1, where it is seen that they tend to lie between 10^{-11} and 3×10^{-9} $M_\odot$ yr^{-1}. These values are consistent with those found by earlier authors, except possibly for γ Cas, for which Hammerschlage-Hensberge et al. (1980) found a value about a factor of 100 higher. This may indicate that his hot Be star does have dominantly SiV in its wind.

The values of $\dot{M}$ in the present study do not correlate well with v sin i, T_{eff}, or M_{bol}. The scatter in these correlations may be due to intrinsic star-to-star variations of some unidentified parameter that controls the winds, to latitudinal dependence of the winds, or to time variations. In the latter regard, it has been shown that in at least one case (δ Cen) the rate of mass loss may vary by as much as a factor of 10 (Snow, Oegerle, and Polidan, 1980), and in two other cases (γ Cas and 59 Cyg) spectacular variations in the ultraviolet CIV and NV profiles have been seen (Henrichs, 1980; Doazan, Kuhi, and Thomas, 1980; Marlborough and Snow, 1980; Doazan et al., this volume) which reflect either changes in $\dot{M}$ or in the wind ionization.

5. THE FeIII SHELL LINES

The narrow FeIII shell lines seen in the ultraviolet spectra of many of these stars (Snow, Peters, and Mathieu, 1979) present a dilemma, for they are usually at or near rest, with velocity widths of about 50 km s^{-1}, in stars that, at the same time, have winds with speeds of several hundred km s^{-1}. There are two possibilities: (1) the shell lines arise outside of the wind zone, in a region where deceleration has occurred; or (2) they form in the wind, at a low level where little or no outwards acceleration has yet occurred.

From the present data it is possible to assess these alternatives somewhat quantitatively. First, the post-deceleration hypothesis (advocated by Thomas, 1980) presumably involves a build-up of material, perhaps in a stationary shock, at some distance from the star, probably quite far out. Snow, Peters, and Mathieu (1979) showed that typical shell column densities are 10^{19}-10^{21} cm^{-2}; at the rates of mass loss found in this study, such column densities would build up in times of order 10 years. Thus, for this hypothesis to work, there must be a loss mechanism capable of dispersing or ionizing FeIII to some other stage, otherwise column densities many orders of magnitude greater than those observed would build-up over the lifetime of a Be star.

To check the alternative hypothesis, the wind column density near zero velocity was computed by integrating the wind optical depth over the appropriate velocity interval (about 50 km s^{-1}), with the result that a shell of order 10^{18} cm^{-2} column density is easily maintained within such an interval of velocity. Thus, the sharp FeIII lines could be produced in the low-velocity portion of the winds, given two assumptions: (1) the ionization of iron changes abruptly from FeIII to high stages at a height where the wind velocity is ∿50-100 km s^{-1}; and (2) the wind in this region is not co-rotating with the star. Both may be reasonable, but both will require further study.

6. SUMMARY

The Be stars in this study have mass-loss rates between 10^{-11} and 3×10^{-9} $M_\odot$ yr^{-1}, if the assumptions regarding the ionization balance of silicon are correct. These values of $\dot{M}$ are 3 to 5 orders of magnitude below those of the most luminous OB supergiants, but otherwise the Be star winds appear similar to those of the more luminous stars, and may simply represent an extension of those radiatively-driven winds to lower luminosities.

The lack of strong correlations with stellar parameters indicates, however, that the Be star winds are dependent on more complex parameters than those of the OB stars, and the complexity may be due to some influential parameter not yet considered, to latitudinal dependence (or other asymmetries) of the winds, or to time variability.

The narrow FeIII shell lines observed in the ultraviolet spectra of many of the Be stars may arise in the low-velocity portions of the winds, below the strong acceleration zone.

Further data are needed, particularly on other ions such as CIV, NV, and especially OVI in order to reduce the uncertainties in the mass-loss rates. It would also be extremely helpful if line profile analysis could be carried out to yield estimates of the stellar inclination angles.

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References

Bohlin, R.C.: 1970, Ap. J., 162, p. 571.

Bruhweiler, F.C., Morgan, T.H., and van der Hucht, K.A.: 1978, Ap. J. Letters, 225, p. L71.

Burton, W.M., and Evans, R.G.: 1976, "IAU Symp. 70, Be and Shell Stars", ed. A. Slettebak (Dordrecht:Reidel), p. 199.

Castor, J.I., and Lamers, H.J.G.L.M.: 1979, Ap. J. Suppl., 39, p. 481.

Code, A.D., Davis, J., Bless, R.C., and Hanbury Brown, R.: 1976, Ap. J., 203, p. 417.

Doazan, V., Kuhi, L.V., and Thomas, R.N.: 1980, Ap. J. Letters, 235, p. L17.

Doazan, V., et al., this volume, p. 415.

Furenlid, I., and Young, A.: 1980, Ap. J. Letters, 240, p. L59.

Hammerschlage-Hensberge, G., et al.: 1980, Astr. Ap., 85, p. 119.

Heap, S.R.: 1976, "IAU Symp. 70, Be and Shell Stars", ed. A. Slettebak (Dordrecht:Reidel), p. 315.

Henrichs, H.: 1980, private communication.

Hutchings, J.B.: 1976, Pub. A.S.P., 85, p. 5.

Lamers, H.J.G.L.M., and Snow, T.P.: 1978, Ap. J., 219, p. 504.

Lamers, H.J.G.L.M., Gathier, R., and Snow, T.P.: 1980, Ap. J. Letters, 242, p. L33.

Lesh, J.R.: 1968, Ap. J. Suppl., 17, p. 371.

Lesh, J.R.: 1972, Astr. Ap. Suppl., 5, p. 129.

Marlborough, J.M., and Snow, T.P.: 1976, "IAU Symp. 70, Be and Shell Stars", ed. A. Slettebak (Dordrecht:Reidel), p. 179.

Marlborough, J.M.: 1977, Ap. J., 216, p. 446.

Marlborough, J.M., and Snow, T.P.: 1980, Ap. J., 235, p. 85.

Peters, G.J.: 1976, "IAU Symp. 70, Be and Shell Stars", ed. A. Slettebak (Dordrecht:Reidel), p. 209.

Peters, G.J.: 1979, Ap. J. Suppl., 39, p. 175.

Slettebak, A., Collins, G.W., Boyce, P.B., White, N.M., and Parkinson, T.D.: 1975, Ap. J. Suppl., 29, p. 137.

Snow, T.P., and Marlborough, J.M.: 1976, Ap. J. Letters, 203, p. L87.

Snow, T.P., Peters, G.J., and Mathieu, R.D.: 1979, Ap. J. Suppl., 39, p. 359.

Snow, T.P., Oegerle, W.R., and Polidan, R.S.: 1980, Ap. J., 242, p. 1077.

Snow, T.P.: 1981, Ap. J., submitted.

Sonneborn, G.H., and Collins, G.W.: 1977, Ap. J., 213, p. 787.

Thomas, R.N.: 1980, private communication.

DISCUSSION

Thomas: 1. I make you a wager that the F_M you list, and the lower limits given by Stalio et al., are low by a factor 100.
2. The Fe II data make me very sceptical that in e.g. γ Cas and 59 Cyg, you can produce it in an atmosphere which is coolest at low radii and hot at large radii.

Snow: 1. I accept.- I'll bet you a Coke that the mass fluxes in the line of sight are not underestimated by a factor of 100 (except for the hottest stars, which may have dominately Si IV, as I have noted).
2. At the moment there is no problem with Fe II, because little or none is seen in any of the stars in my list. I agree that this might be a difficult thing to explain, if we find stars with Fe II shell lines, particularly the visible-wavelengths ones (because they require high densities), but even then it may be possible to have Fe II close to the star. I remind you of my comment that charge-exchange reactions can produce Fe II from Fe III very efficiently.

Doazan: In many shell spectra one observes in the UV, in addition to Fe III lines shell lines of Fe II. These lines cannot be formed near the photosphere. I see no other possibility than to form them far in the outer atmosphere. In this case, the observed low velocity imposes a velocity law which is decelerated in the outer regions and which is completely different from the one you assumed.

Snow: Let me elaborate on my answer to a similar question from Dick Thomas. It is possible to have some Fe II near the star, if even a little He I is present to form it through charge- exchange with Fe III. Even without this, a little Fe II can survive the ionizing flux of the star, although I admit it would be very little for the hotter Be stars. However, I must stress that none of the stars in my sample have notable Fe II shell lines, so at the moment there is no inconsistency to worry about. I know that one of the stars you are referring to is 88 Her, which is cooler than my sample stars, so the presence of Fe II close to the star, may be even less difficult to explain.

Persi: I agree with Dr. Snow that to determine the mass-loss rate from IR observations the terminal velocities derived from UV observations cannot be used, because IR and UV radiations originate in different regions of the envelope. In any case the ratio $\dot{M}/v_\infty$ derived for γ Cas from our IR observations is greater than derived from UV observations by a factor $>$ 100.

Snow: This discrepancy bears on a central issue that has come up many times during this symposium, and perhaps is most obvious in this case. There must be regions with very different physical conditions in the circumstellar environment, so we must be very careful when comparing data obtained in different ways in different wavelength regions.

Stalio: I am glad to see that you have used the same method to determine the wind components as we did; our ways of computing mass-loss rates are, however, different. Because of the many uncertainties that arise and you have mentioned, especially on ionization equilibrium, we made the choice of computing lower limits. I have the impression that you made the same at least for some stars that we have in common.
Concerning the correlation with v sin i it would be interesting to put together our results.