THE HD 5980 MULTIPLE SYSTEM: MASSES AND EVOLUTIONARY STATUS

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ABSTRACT

New spectroscopic observations of the LBV/WR multiple system HD 5980 in the Small Magellanic Cloud are used to address the question of the masses and evolutionary status of the two very luminous stars in the 19.3 day eclipsing binary system. Two distinct components of the N v 4944 Å line are detected in emission and their radial velocity variations are used to derive masses of 61 and 66 M☉under the assumption that binary interaction effects on this atomic transition are negligible. We propose that this binary system is the product of quasi-chemically homogeneous evolution with little or no mass transfer. Thus, both of these binary stars may be candidates for gamma-ray burst progenitors or even pair instability supernovae. Analysis of the photospheric absorption lines belonging to the third-light object in the system confirm that it consists of an O-type star in a 96.56 day eccentric orbit (e = 0.82) around an unseen companion. The 5:1 period ratio and high eccentricities of the two binaries suggest that they may constitute a hierarchical quadruple system.

Key words: binaries: eclipsing – stars: evolution – stars: individual (HD 5980) – stars: massive – stars: Wolf−Rayet

Online-only material: color figures

1. INTRODUCTION

Massive stars play an important role in the chemical evolution of galaxies and in shaping the interstellar medium. They are found in regions of recent star formation and near the Galactic center, and they are associated with fascinating phenomena such as the formation of neutron stars and black holes. But despite their importance, there are still gaps in our understanding of their structure and evolution. One of these gaps involves the manner in which they reach the final stages of evolution and explode as supernovae (SNe). In particular, there is great uncertainty concerning the amount of mass that is lost in the post-main sequence evolutionary stages, thus leading to uncertainties in the stellar structure prior to the final SN event. Central to this issue is the role played by the Luminous Blue Variables (LBV).

According to current evolutionary scenarios, classical LBVs are stars in transition from the main sequence to the Wolf−Rayet (WR) stage and they are in the process of shedding their outer layers to expose the products of the CNO nuclear reactions taking place in their centers (Maeder & Meynet 1987; Langer et al. 1994, Langer 2012). Enhanced mass-loss rates are observed to occur during the maxima of the variability cycles when up to \( \sim 10 M_\odot \) can be ejected, as illustrated by \( \eta \) Carinae’s mid-19th century event (Humphreys & Davidson 1994). However, the underlying processes that cause the LBV instability are unknown (see, for example, Vink et al 2012). Thus, a crucial ingredient in the models of massive star evolution is missing. This problem is compounded by the fact that the current mass of the LBVs is in general poorly constrained.

The Small Magellanic Cloud (SMC) harbors an LBV eclipsing binary that has been studied since the 1950s and which underwent a major eruption in 1994. It has the highest luminosity of all the well-studied LBVs, except perhaps for \( \eta \) Carinae, and also shares with \( \eta \) Car the exceptional characteristic of showing a WR type spectrum. Furthermore, both systems are associated with massive and young stellar clusters, HD 5980 lying in the periphery of NGC 346. Its wind properties and many of its fundamental parameters have been derived in a series of studies based on IUE, HST/STIS, and optical observations (Perrier et al. 2009; Georgiev et al. 2011; Foellmi et al. 2008; Koenigsberger 2004; Koenigsberger et al. 2010). The mass determinations so far, however, have a large degree of uncertainty, a problem we address in this investigation. In addition, the long baseline of data now available allows a re-assessment of the nature of the third component of the system, which we now can confirm consists of a binary system.

The organization of this paper is as follows. In Section 2 we review the characteristics of the HD 5980 system; in Sections 3 and 4 we describe, respectively, the observational data and the radial velocity (RV) measurements; Section 5 presents the orbital solutions; in Section 6 we discuss the evolutionary paths that may explain HD 5980’s current state; and in Section 7 we summarize the conclusions.

2. DESCRIPTION OF THE HD 5980 SYSTEM

The spectrum of HD 5980 is the sum of several components. Most prominent are the two emission-line stars belonging to the close eclipsing binary system \( (P_{AB} = 19.3 \text{ days}) \). Following the
convention introduced by Barbá et al. (1996), the star “in front” at orbital phase $\phi = 0.00$ is labeled Star A and its companion is labeled Star B. In addition, there is a third component, Star C, whose photospheric absorptions remain stationary on the 19.3 day orbital timescale and which contributes a significant fraction of the observed light (Breysacher & Perrier 1991; Niemela 1988; Koenigsberger et al. 2002; Perrier et al. 2009).

HD 5980’s emission-line spectrum is highly variable on orbital and decade-long timescales. The short-timescale variability is produced by the periodic Doppler motion of the emission lines that are formed in the winds of Stars A and B combined with binary interaction effects, among which the ones that are most likely dominant are irradiation, wind–wind collisions (WWCs) and wind eclipses. Different lines are affected to a different degree, depending on their transition probabilities and formation mechanisms. The long-term variation is due to Star A’s changing wind structure, which appears to be triggered by variations in the hydrostatic radius (Koenigsberger et al. 1998; and Georgiev et al. 2011).

Star B is thought to be a WN4 star (Breysacher et al. 1982; Niemela 1988), and hence it has a stellar wind that contributes significantly to the emission-line spectrum. Traces of its broad UV emission lines can be seen underlying the narrow emissions that appeared during the 1994 eruptive event (see Koenigsberger 2004, Figure 11). Furthermore, the shape of the eclipse light curve observed in the late 1970s leads to the conclusion that Star B possesses an extended occulting region having a radius of 0.269$a$, where $a$ is the semi-major axis of the orbit (Perrier et al. 2009). Both of these facts lend credence supporting the view that Star B is a WNE-type star.

The spectrum of Star C is that of an O-type supergiant. Its optical spectrum is dominated by H-Balmer, He I, and other low-ionization photospheric absorption lines, while in the UV it presents a modest P Cygni Si iv 1400 Å profile (Georgiev et al. 2011). The RV of the O III 5592 Å photospheric line was found to display a periodic behavior with $P_C \sim 97$ days, and the solution to this RV curve yielded a highly eccentric $e \sim 0.8$ orbit, leading to the conclusion that Star C is also a binary system (Schweickhardt 2000).

The determination of the masses of the individual components in the system is a challenging problem due to the very broad and variable emission lines. Using photographic spectra obtained during 1981–1983, Niemela (1988) estimated masses $M_A \sim 46 M_\odot$ and $M_B \sim 43 M_\odot$ for Star A and B, respectively. Foellmi et al. (2008) analyzed high dispersion spectra obtained in 1998, 1999, and 2005 and concluded that, if the observed line profiles can be interpreted as being formed primarily in the winds of Star A and Star B (as opposed to in the WWC region; see Section 5.3), then the measured RV variations implied $M_A \sim 58–79 M_\odot$ and $M_B \sim 51–67 M_\odot$. The stellar wind analysis performed by Georgiev et al. (2011) led to the conclusion that for a stable solution connecting the photosphere with the stellar wind to be possible, $M_A \geq 90 M_\odot$. Moffat et al. (1998) assumed that at least part of the photospheric absorptions in the spectrum arise in Star A, from which they derived masses $M_A \sim 28–38 M_\odot$ and $M_B \sim 6–18 M_\odot$. A possible explanation for the modulation on the 19.3 day period of the absorption lines claimed by Moffat et al. (1998) was put forth by Georgiev & Koenigsberger (2004). They showed that if Star A is a rapid rotator, the superposition of its broad and shallow photospheric lines could produce a slight RV fluctuation in Star C’s sharper lines and lead to the false conclusion that they follow the 19.3 day orbital period, albeit with a very small RV curve amplitude.

### 3. OBSERVATIONAL MATERIAL

The observations of HD 5980 analyzed in this paper were obtained at Las Campanas Observatory and consist of seven low dispersion spectra obtained with the 6.5 m Magellan telescopes in 2009 and 2010; 25 echelle spectra obtained with the du Pont telescope between 2006 and 2013; and six spectra obtained on the 6.5 Clay Magellan telescope between 2007 and 2012. Tables 1 and 2 list the year the spectrum was obtained, the telescope and instruments used, Heliocentric Julian Date of observation, and orbital phase for $P = 19.2654$ days and $T_0 =$ HJD 2,443,158.705 (Sterken & Breysacher 1997).

The Baade 6.5 m (Magellan-I) low dispersion spectra were obtained with IMACS in its long camera mode (4) with a 600 line mm$^{-1}$ grating that provided spectral coverage from...
roughly 3700 to 6800 Å at a spectral resolution of 1.4–1.5 Å ($R \sim 4000$). One of these spectra was described in Koenigsberger et al. (2010). The Clay 6.5 m (Magellan-II) low dispersion spectra were obtained with the Magellan Echellette (MagE) using the 1″ slit providing a spectral resolution of 1 Å. Thirteen echelle orders were extracted covering the wavelength region from 3130 Å to 9400 Å. The signal-to-noise ratio (S/N) ranges from 100 to 200 for a single 150 s exposure. The usual ThAr comparison lamp was used for wavelength calibration, and the data were reduced using an adaptation of the mtools package, available for the reduction of Magellan Inamori Kyocera Echelle (MIKE) spectra (see below) in combination with IRAF8 echelle routines for the reduction of MIKE spectra.9 Spectra of spectrophotometric standard stars observed during the same nights of observation were used to derive a sensitivity function. The individual flux calibrated echelle orders were then normalized and merged in the final spectrum. One of these MagE observations was already described in Georgiev et al. (2011).

The high resolution du Pont telescope spectra were obtained with the echelle spectrograph and a 1″ slit. The spectral resolution of these data ranges from 0.15 to 0.22 Å ($R \sim 25,000$) and the wavelength coverage goes from 3500 Å to 8500 Å.

The high resolution Clay (Magellan-II) spectra were obtained with the MIKE spectrograph using a 0.7′′ slit and applying a 2 × 2 binning to both blue and red detectors. This configuration results in a spectral resolution of ~34,000 (FWHM ranging from 0.10 to 0.25 Å). Reductions were carried out with the specially designed IRAF scripts contained in the mtools package developed by Jack Baldwin and available at the Las Campanas web site. The reduced spectra were flux calibrated using spectra of spectrophotometric standards obtained during the same observing nights.

In addition to the new spectra described above, we used data obtained in 1998, 1999, 2005, and 2006 with FEROS on the 2.2 m MPI telescope at La Silla Observatory, previously described in Kaufer et al. (2002), Schweichhardt (2000), and Foellmi et al. (2008).

4. RADIAL VELOCITIES AND ORBITAL SOLUTIONS

In the following sections we will describe the procedure followed for obtaining the RV curves of the two emission-line objects and of the star responsible for the absorption lines. For this, it is useful to keep in mind that the eclipses of the 19.3 day binary occur at $\phi = 0.00$ (Star A “in front”) and $\phi = 0.36$ (Star B “in front”) and maximum receding velocity of Star A is at $\phi \sim 0.1$.

4.1. Emission Lines

Figure 1 shows the spectra obtained in 2009 at orbital phases $\phi = 0.038$ and 0.944 in the $\lambda \lambda 3100–7400$ wavelength region. The dominant features are emission lines of He II, N IV, N v, and C IV; He I is weakly present in emission at $\lambda 5875$. All the major emission lines undergo prominent shape changes over the orbital cycle, similar to those found in previous optical data sets (Breysacher et al. 1982; Breysacher & François 2000; Moffat et al. 1998; Foellmi et al. 2008; Koenigsberger et al. 2010). The dominant type of variability involves the FWHM intensity which systematically is greater at elongations than at eclipses. Star A and Star B are believed to possess stellar winds of similar ionization characteristics and the orbital motion of these stars combined with the interactions and the eclipses produces the FWHM variations. In Section 5.3 we discuss the validity of this assumption in the light of the various interaction effects that are present in the Star A + Star B system, and we argue in favor of the possibility of using the higher ionization emission lines for deriving the orbital parameters. Because the ionization rapidly decreases outward, the bulk of their emission arises in the innermost regions of the wind and are thus less sensitive to external perturbations.

Narrow lines are also the best ones to measure on the echelle spectra, where uncertainties in the rectification of the individual orders represents a challenge for accurate measurement of very broad and variable emissions. In addition, it is advisable to select lines that lie near the center of the order where the S/N is greatest and that are not blended with other lines. We found that the line that is best suited for the RV curve analysis is N v $\lambda 4944$. This is one of the three N v lines that lie in the same echelle order at $\lambda \lambda 4944.37, 4932.00$, and 4952.07. The first of these lies closest to the center of the order in most of our spectra and it is the only one of these three lines that is clearly visible. This is consistent with its intensity in our Star A CMFGEN model, where it has an equivalent width ~8 times greater than that of $\lambda 4932$ and 4952.

Figure 2 displays the rectified10 echelle orders containing these N v transitions and stacked from bottom to top with increasing orbital phase plotted on a velocity scale corrected for an assumed SMC systemic velocity of +150 km s$^{-1}$. It shows that N v $\lambda 4944$ splits into two components during the phase intervals: 0.09–0.24 (when Star A is receding from the observer) and 0.58–0.81 (when it is approaching). A similar pattern is

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8 IRAF is written and supported by the National Optical Astronomy Observatories (NOAO) in Tucson, Arizona. NOAO is operated by the Association of Universities for Research in Astronomy (AURA), Inc. under cooperative agreement with the National Science Foundation.

9 mtools package: http://www.lco.cl/telescopes-information/magellan/instruments/mike/iraf-tools.

10 The echelle orders were rectified by tracing the continuum level on each order and then using the resulting curve for the normalization.
also observed to occur in the FEROS data of 2005–2006, as illustrated in Figure 3.

The RVs of N v 4944 were measured with a two-Gaussian deblending routine in IRAF. These measurements were first performed on the non-rectified LCO echelle orders and then repeated on the rectified orders. In the spectra in which the two components are well separated, the largest uncertainty is $\sim 10$ km s$^{-1}$ and stems from the definition of the continuum level. At eclipse phases where only one centrally located emission is visible, the measured RV is assigned to the uneclipsed component, with the uncertainty stemming also from the continuum level. When the lines are not well separated, the uncertainty can be as large as $\sim 30$ km s$^{-1}$.

Consistent measurements were not possible on a number of spectra due in general to a combination of too poor S/N, spurious features, and cosmic ray hits, and these spectra were excluded from the N v 4944 analysis. The 2005–2006 FEROS exposures were performed in pairs, and each individual spectrum was measured and the results averaged. The results are presented in Table 3 and Figure 4.

We searched for other N v lines that could be measured in a similar fashion as N v 4944, but they are all either too weak or lie too close to the edges of the orders, except for $\lambda \lambda$ 4603–19 which, however, possess strong overlapping P Cygni absorption components which do not allow reliable measurements.

Another line that is amenable for measurement is N iv 4057 Å. It was measured by Foeckl et al. (2008) in the 1999 and 2005 FEROS spectra by assuming that it consists of two components that are Doppler-shifted due to orbital motion. We have applied the same procedure to measure this line in the LCO spectra listed in Table 2. To do so, we fixed the limits for the deblending routine at $\pm 500$ km s$^{-1}$, which therefore neglects the broad emission-
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Figure 4. Radial velocities of the N\textsc{v} 4944 Å emission line obtained in 1999–2008 (triangles), 2009–2012 (squares), and 2013 (crosses) plotted as a function of the 19.2654 day orbital phase. Error bars indicate the measurement uncertainties. The dash curves correspond to the orbital solution listed in Table 6 with a systematic shift of $-20$ km s\(^{-1}\).

(A color version of this figure is available in the online journal.)

Figure 5. Radial velocity measurements of the N\textsc{iv} 4057 emission line from the rectified LCO orders and from the archival FEROS data plotted as a function of the 19.2654 day orbital phase. Open and filled-in symbols correspond, respectively, to Star A and Star B. The larger dispersion here compared to the plot of N\textsc{v} 4944 Å is most likely related to stronger interaction effects in this N\textsc{iv} emission line.

(A color version of this figure is available in the online journal.)

Figure 6. Selected spectral regions showing photospheric absorption from Star C in the low dispersion spectra obtained at orbital phases $\phi_{AB} = 0.038$ (lower tracing) and $0.944$ (upper tracing; displaced vertically for clarity). Shown are He\textsc{i} 4387 and 4471 Å (left panel) and O\textsc{iii} 5592, C\textsc{iv} 5801, 5812, and He\textsc{i} 5875 Å (right panel).

(A color version of this figure is available in the online journal.)

Curves of N\textsc{iv} 4057 Å at orbital phases 0.25–0.45, most likely a consequence of epoch-dependent variations in Star A’s wind structure.\(^{11}\)

4.2. Photospheric Absorption Lines

It is now well-established that the sharp photospheric absorptions arise in Star C, but the presence of such absorptions arising in Star A and/or Star B cannot be excluded. A search for photospheric absorptions has repeatedly been conducted (Niemela 1988; Koenigsberger et al. 2002) with no success. This can be explained if the optical depth of the stellar wind is too large for the spectrum of the hydrostatic radius to be visible, an explanation that is supported by the CMFGEN model which indicates that the optical depth at the sonic radius is $\sim 1.8$ (Georgiev et al. 2011; model fit to the 2009 spectra). However, there is evidence that the wind is becoming weaker so that photospheric absorptions may be expected to become visible, prompting a new search in the current data set.

Ten of the du Pont Telescope observations were obtained over a 32 night timespan in 2010 and cover the orbital phases 0.58–1.24. We first searched these spectra for photospheric absorption lines with RV variations that would allow them to be associated with either Star A or Star B. Although many potential identifications were made, we were unable to convincingly find any one feature that appears in several spectra with RV shifts that would prove that it arises in one of these stars. The search was then extended to the other spectra with similar inconclusive results. It is worth noting that photospheric absorptions in a rapidly rotating star are difficult to detect in general and more so if they are superposed on broad and variable emission lines as might be the case for Star A and/or Star B.

We now shift our focus to the photospheric absorptions that are visible in the spectra and do not partake in the 19.3 day orbit. These include the H-Balmer series out to H\textsc{ii}, as well as prominent He\textsc{i} 4471, 5875 Å, O\textsc{iii} 5592 Å, and C\textsc{iv} 5801, 5812 Å, and weaker absorption attributable to He\textsc{ii} 4542 and 5411 Å. The absorptions are most clearly visible at orbital phases when the underlying emissions are broadest (such as at $\phi = 0.944$ in Figure 6). As mentioned previously, the O\textsc{iii} 5592 Å line was shown to undergo RV variations consistent with a $P_C \sim 96.5$ days, $e \sim 0.8$ orbit (Schweickhardt 2000).

\(^{11}\) In this phase interval Star B is “in front” of the companion, and changes in the wind structure of Star A would affect the geometry of the wind–wind collision region; however, given the larger Star A wind density, the irradiation effects would also be more important and back-scattered emission would contribute to the “blue” emission-line wing.
In addition, we measured the C
measurement, all of which are relatively uncontaminated by neigh-
Figure 8 displays these values as a function of phase computed
the lines in the LCO and the FEROS spectra, respectively, and
and 5 list the RVs obtained by fitting Gaussians to the cores of

conclusion that
photospheric absorption RVs and convincingly showed that it was not present.

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13 Schweickhardt searched for modulation on the

HJD Phase Phase
1094.797 0.935 0.082 251 183
1100.773 0.245 0.144 214 208 189
1133.691 0.954 0.485 205 192 132
1138.512 0.204 0.535 166 141 128
1145.547 0.569 0.608 198 174
1150.508 0.827 0.659 181 201 61
1174.523 0.073 0.908 126 91
1176.555 0.179 0.929 82 108 87
1181.516 0.436 0.980 80 174
1183.523 0.541 0.001 122: 111: 154
1185.562 0.646 0.022 229 94:
1187.523 0.748 0.043 234 236 194
1189.516 0.852 0.063 232:
1191.516 0.955 0.084 268:
1193.531 0.060 0.105 217 216:
1197.535 0.268 0.146:
1374.918 0.475 0.983 68 69 44
1375.922 0.527 0.994 88 88 70
1379.902 0.731 0.917
1381.902 0.838 0.056 212 219:
1382.852 0.887 0.066 200 216 260
1383.855 0.939 0.076 186 222 245
1384.809 0.989 0.086 195 240 208
1386.891 0.097 0.107 180 216:
1388.840 0.198 0.128 172 239:
1389.883 0.252 0.138 184 216 216
1390.828 0.301 0.148 178 207 224
1392.828 0.405 0.169 182 204 203
1393.871 0.459 0.180 156 198 131
3538.894 0.799 0.394 164 179 132
3541.890 0.955 0.425 174 179 143
3561.877 0.993 0.632 161 167 177
3576.795 0.768 0.787 146 143 139
3635.520 0.815 0.395 176 183 167
3638.763 0.984 0.428 162 185 296
3641.662 0.134 0.458 169 167 131
3665.527 0.373 0.706 151 162 174
3715.789 0.982 0.226 211 202 204
3716.539 0.021 0.234 186 182 183
3731.549 0.799 0.389 192 179 173
3735.551 0.007 0.431 184 183 180
3738.539 0.163 0.462 170 177 181

are clustered by chance in the two phase intervals \( \phi_C = 0.0–0.3 \) and 0.7–1.0, which correspond, respectively, to times of

maximum and minimum RVs. This is shown by plotting with
different symbols the data in each of these two phase intervals
in the top left panel of Figure 8. This serves as a cautionary
tale to be heeded when finding modulations in data sets that
incompletely sample the actual period. Furthermore, if we now
plot the LCO data from observations prior to 2010 together with
the 1998–1999 data, the plot becomes a scatter diagram as
found by Schweickhardt (2000) for the 1998–1999 observations.
Hence, we conclude that \( P_C \) does indeed exist and can only
be associated with \( Star\ C \) unless \( Star\ A \) or \( Star\ B \) undergo
some kind of pulsation cycle on the \( P_C \) timescale which also
masks the RV variations on \( P_AB \), something we do not consider
very likely.

Little can be said at this time regarding the nature of
the secondary component of the \( Star\ C \) system. There is marginal
evidence for the presence of weak absorptions that may move

12 The 1998–1999 data are RVs measured on the same spectra that were used
by Schweickhardt (2000).
13 Schweickhardt searched for modulation on the \( P_AB \) period in the \( O\ iii 5592 \)
photospheric absorption RVs and convincingly showed that it was not present.
Figure 7. RV measurements of the Star C lines plotted against phase computed with \( P_C = 96.56 \) days. Open triangles (LCO spectra) and crosses (FEROS spectra) correspond to the photospheric absorptions of O\textsc{iii} and He\textsc{i}; open squares (LCO) and filled-in triangles (FEROS) correspond to the weak C\textsc{iii} 5696 emission line.

(A color version of this figure is available in the online journal.)

Figure 8. RVs of the photospheric absorption lines O\textsc{iii} 5592 and He\textsc{i} 4471, 5875 Å plotted as a function of phase computed with \( P = 19.2654 \) days (top) and 96.56 days (bottom) for different epochs of observation. The apparent modulation of the 2010–2012 data on the 19.2654 day period is due to a chance coincidence: open triangles and filled squares indicate, respectively, data with \( \varphi_C = 0.0–0.3 \) and \( 0.7–1.0 \), phase ranges at which maximum and minimum of the 96.56 day modulation occur.

(A color version of this figure is available in the online journal.)

Finally, we note that the RVs of the weak C\textsc{iii} 5696 Å emission line follow the same trend as the photospheric absorptions and thus it originates in Star C.

5. ORBITAL SOLUTIONS

5.1. The 19.3 Day WR/LBV System

We employed the FOTEL\textsuperscript{14} program (Hadrava 2004) to fit the RV variations of the two Doppler-shifted emission lines of N\textsc{v} 4944. Assigning the same weight to all the data points and fixing \( i = 86^\circ \) from Perrier et al. (2009), the best solution gives \( M_A = 61 \pm 10 M_\odot \), \( M_B = 66 \pm 10 M_\odot \), where the uncertainties refer to the goodness of fit. Our value for the eccentricity \( e = 0.27 \pm 0.02 \) is slightly smaller than that derived by Perrier et al. (2009) from the solution of the light curve (\( e_{\text{Perrier}} = 0.314 \pm 0.007 \)) but coincides with that obtained by Schweickhardt (2000) from the RV curve solution of the 1998–1999 FEROS data, \( e = 0.284 \). Our argument of periastron \( \omega = 134^\circ \pm 4 \) coincides within the uncertainties with those of Perrier et al. (2009) and Schweickhardt (2000).

The FOTEL fit yields a period \( P = 19.2656 \pm 0.0008 \) days which coincides within the uncertainties with that derived by Sterken & Breysacher (1997), \( P_{AB} = 19.2654 \) days. Hence, we continue to adopt the Sterken & Breysacher (1997) ephemeris in this paper. Our results do not change if the period is fixed to this value. The time of periastron that is found by our fit, HJD2,451,424.97 \( \pm 0.25 \), corresponds to orbital phase \( \varphi_{\text{peri}} = 0.075 \).

\textsuperscript{14} http://www.asu.cas.cz/~had/paifotel.pdf

in anti-phase with the primary lines as illustrated in Figure 9. Significantly higher S/N spectra are required for further analysis of this issue, however.
The parameters for the Star A and Star B system are listed in Table 6. The predicted RV curves are plotted in Figure 4, displaced by \(-20 \text{ km s}^{-1}\), consistent with an expected systematic blueshift in NV 4944 Å due to occultation by the stellar core of the wind that is receding from the observer. Our mass determination is consistent with the estimate derived in Foellmi et al. (2008), but it is larger than that of Niemela (1988) and our \(M_1\) is significantly smaller than the 90 \(M_\odot\) estimated by Georgiev et al. (2011).

Niemela (1988) used the RV curves of the emission lines NV 4603 Å and NV 4058 Å assumed to arise, respectively, from Star A and Star B. This explains the somewhat smaller masses because the NV line is affected by the P Cygni absorption component of the neighboring NV 4620 Å line while the NV emission contains a contribution also from Star A.

An answer to the question of whether the 90 \(M_\odot\) deduced by Georgiev et al. (2011) can be reconciled with our current results is less straightforward. The spectroscopic mass in the CMFGEN model is based on the estimated luminosity and the proximity of the star to the Eddington limit. For the 2009 model, Star A has \(L/\text{Edd} = 0.50\). If we were to use 0.6 as an upper limit for the Eddington ratio, we would obtain \(M_1 = 75 \ M_\odot\), in better agreement with that deduced from the RV measurements. With \(L/\text{Edd} \geq 0.60\) it becomes difficult to construct a photospheric model since the radiation force will potentially exceed gravity below the photosphere because of the additional opacity due to iron (as we will discuss below; see Figure 14). However, in this regime there are radiation instabilities that need to be addressed, and clumping can potentially reduce the effective opacity (e.g., Shaviv 2001a, 2001b).

Part of the discrepancy could also arise from the deduced luminosity since the deduced spectroscopic mass scales with \(L\). If the error in \(T_\text{e}\) is \(\sim 3\%\) (2000 K for \(T_\text{e} = 60,000\) K) the error in \(L\) is approximately 10\%. Given the additional uncertainties arising from the spectroscopic decomposition an error in \(L\) of 20\% is not unreasonable, and this translates to an error of almost 20 \(M_\odot\).

Finally, we note that the uncertainty in the mass determinations quoted in Table 6 does not include the uncertainties derived from possible contributions to the NV 4944 Å emission lines arising in, for example, the WWC region, a point that will be discussed in Section 5.3.

### 5.2. The 96.56 Day Binary System

The RVs of the O III 5592 and He I 4471, 5875 Å photospheric absorptions were also analyzed with the FOTEL program (Hadrava 2004), from which the following orbital elements were derived: \(P = 96.56\) days, \(e = 0.82, \omega = 252^\circ\), and a time of periastron HJD2,451,183.40. As shown in Table 7, these values are consistent with those obtained by Schweickhardt (2000).

The systemic velocity that is derived from the average of the different data sets (LCO, FEROS, three different lines) is \(V_0 = 164 \pm 5\) km s\(^{-1}\). This is in contrast with the value we find for Stars A and B, \(V_0^{AB} = 131 \pm 3\) km s\(^{-1}\). However, as pointed out above, a systematic blue shift in the NV 4944 Å emission line is expected, so once the \(-20\) km s\(^{-1}\) estimated shift is contemplated, the value of \(V_0^{AB}\) is very similar to that of \(V_0\), consistent with both binary systems lying at the same location within the SMC.

### 5.3. Do the Emission Lines Reflect Orbital Motion?

The basic assumption that is made when using spectral lines to determine the orbital parameters of a binary star is that these lines have a constant shape over the orbital cycle. In the case of an emission-line star, this translates into the assumption that the distribution of emitting atoms in the material surrounding the star is symmetric and remains constant over the orbital cycle as viewed in the frame of the observer. If these conditions are not met, then the shape of the line profiles will change over orbital phase, thus introducing a systematic deviation in the RV measurements with respect to those that yield orbital motion.

The issue of where in the HD 5980 system the emission lines are formed has been a long-standing topic of debate and has limited the efforts to determine the masses of Star A and Star B. Moffat et al. (1998) and Breysacher & François (2000) speculated that the bulk of the emission lines arises not in the stellar winds, but in the shock cone region where the two winds collide (henceforth referred to as the WWC). If this were the case, then the RV variations would describe the velocity structure along the shock cone walls, projected along the line of sight to the observer, and their use for deducing the stellar masses would be very limited.

Although the WWC shock walls may indeed give rise to emission lines, there are additional effects that may also distort the emission-line profiles. For example, the wind region belonging to the star with the weaker wind that is enclosed by the WWC constitutes a “hole” in the spherical distribution of the stronger wind star. It has been shown that this effect results in weaker emission at particular wavelength intervals that change with orbital phase (Georgiev & Koenigsberger 2004). Another interaction effect is the irradiation of Star A’s wind by Star B, which may alter the ionization and velocity structure (Gayley et al. 1997) on the affected hemisphere. One-dimensional model calculations of the P V line profile variations in FUSE spectra led to the conclusion that the wind structure of Star A is different on the hemisphere facing the companion from that of the opposite hemisphere (Koenigsberger et al. 2006), supporting this scenario.
Figure 10. Shown (red, solid) is the radial contribution to the flux ($N\nu 4994$ Å line and underlying continuum) measured over a band extending from 4942.3 to 4945.3 Å. Also shown (blue, broken) is the radial contribution to the continuum flux near the $N\nu 4994$ line. These data were obtained from the CMFGEN model constructed for Star A for the 2009 epoch (Georgiev et al. 2011). The radius variable is given in units of $R_{10} = 19.3 R_\odot$, which corresponds to the location where the continuum optical depth $\tau_{Ross} = 10$. Adopting the orbital elements listed in Table 6, the separation between Stars A and B at periastron is $\log(R_{peri}/R_{10}) = 0.76$ (indicated by vertical line), which is outside of the wind region where this $N\nu$ line primarily forms.

(A color version of this figure is available in the online journal.)

Although a proper analysis of irradiation, WWC, and asymmetric wind structures requires the use of two- or three-dimensional non-LTE radiative transfer computations, the dominant effect of these interaction effects is expected to be present in the outer wind regions. Hence, as we shall argue below, emission lines that are formed in the innermost wind regions, such as $N\nu 4944$ Å, are likely to be much less affected.

Weak emission lines such as $N\nu 4944$ Å are formed in a small volume, one that is comparable to that of the continuum forming region, as we illustrate in Figure 10 which shows the extent of the $N\nu 4944$ Å forming region computed by the CMFGEN model for Star A (Georgiev et al. 2011). The bulk of this emission arises at $\log(r/R_{10}) \leq 0.3$, where $R_{10}$ is the radius at which the continuum optical depth $\tau$ is 10. Adopting $R_{10} = 19.3 R_\odot$ from Georgiev et al. (2011), we find that the bulk of the $N\nu 4944$ Å is formed at $<40 R_\odot$, which is significantly smaller than the orbital separation, even at periastron passage ($r_{peri} \sim 110 R_\odot$). This supports the assumption that the RV curve of this $N\nu$ line is likely to be unperturbed by interaction effects.

A further test is to compare the CMFGEN model line profiles with the observations, for which we synthesized spectra of the HD 5980 system under the following conditions: (1) the relative continuum brightnesses of Star A, B, and C in the visual spectral region are, respectively, 40%, 30%, and 30%; (2) the emission-line spectrum of Star B is approximated with the CMFGEN emission-line spectrum of Star A computed by Georgiev et al. (2011) for the 2009 epoch, scaled to the relative luminosity of Star B; (3) the spectrum of Star C is represented by the appropriate CMFGEN model (Georgiev et al. 2011).

The CMFGEN model spectra for Star A and Star B were Doppler-shifted to the RVs measured on the $N\nu 4944$ Å line in the FEROS spectra at orbital phases $\phi = 0.162$ and 0.799. These were then combined with each other and with the spectrum of Star C to create the synthetic spectrum of the triple system at these two orbital phases.

Figures 11 and 12 present a selection of emission-line profiles in the spectra obtained in the year 2006 at orbital phases $\phi = 0.162$ (left panel) and $\phi = 0.799$ (right panel) compared with the synthetic spectrum computed for the corresponding phases. Other than the discrepancies in line strengths which are to be
expected given the uncertainties in the input parameters for the CMFGEN model, the synthetic triple-star spectrum reproduces quite well the observations. Hence, there is no indication that the lines shown in these figures are affected in any significant way by the interaction effects in the binary system.

It is also important to note that the N\textsc{v} 4944 Å emission at orbital phase $\phi = 0.162$ consists of two separate lines, both of which reach the local continuum level; i.e., they are not really “double peaked” emission lines, a shape predicted from models of emission arising in the WWC shock cone walls (see, for example, Henley et al. 2003).

Finally, we note that although winds with similar momentum ratios lead to a WWC region in which the phase-dependence of emission line profiles is similar to that of two distinct wind sources, to the extent of our knowledge, no calculation of the intensity of such emission lines has been performed. Thus, it is not possible to assess the relative importance of this emission compared to that which arises from the inner wind regions of the two stars.

Hence, we conclude that, given the information for HD 5980 that is available at this time and given the lack of models that predict the strength of the emission lines expected to arise in a WWC region, the assumption that the N\textsc{v} 4944 Å emission line arises primarily in the inner stellar winds of the stars is a sound one.

Clearly, however, this conclusion is based on the premise that Star B possesses an emission-line spectrum that arises in a high-ionization stellar wind typical of WNE-type stars. This premise is based on observations obtained prior to and during the eruptive state (Breyasaki et al. 1982; Niemela 1988; Koenigsberger 2004). If Star B were to be a non-WR star, then the WNE-type spectrum that has been observed for $>60$ yr would have to arise entirely in Star A and, potentially, in the WWC region, and the conclusions of this paper would have to be re-assessed. It is important to note, however, that the wind momentum ratio under these circumstances would most likely be such that the WWC shock cone would be folded around Star B and hence, a highly blueshifted emission component would be observed in lines of He\textsc{i} and He\textsc{ii} near the conjunction when this star is between us and its companion,$^{15}$ something that is not observed.

6. EVOLUTIONARY STATUS

In this section we will examine the evolutionary scenarios that lead to a system such as the 19.3 day binary in HD 5980.

A preliminary examination (Koenigsberger 2004) indicated that the luminosity and effective temperature of Star A place it on the Hertzsprung–Russell diagram (HRD) along the evolutionary track of a 120 $M_\odot$ zero-age main sequence (ZAMS) star at an age of $\sim$3 Myr. This is very close to the age of the NGC 346 cluster (Mokiem et al. 2006), which would indicate that such a massive star is near the final evolutionary phases leading up to an SN event. A similar result is reached if we use the most recently available grid of models for the SMC metallicity (Georgy et al. 2013). Specifically, Star A’s luminosity corresponds to that of a non-rotating 120 $M_\odot$ ZAMS model near the end of core helium burning at an age of 2.952 Myr or to that of a 85 $M_\odot$ model with initial rotation velocity of 400 km s$^{-1}$ at an age of 3.63 Myr. However, both of these models predict a large increase in radius prior to these ages. Since the current Roche–lobe radius is only $\sim$57 $R_\odot$, the interaction with the companion would have modified the subsequent evolution from that predicted by the single-star models. Thus, HD 5980 needs to be analyzed within the context of binary evolution or within scenarios in which the large radius increase in post-main sequence stages can be avoided.

The classical binary scenarios involve the assumption that a star evolves as a single star until its radius reaches some critical value, assumed to be the Roche radius. At this point in the evolution, the nature of the close companion becomes a crucial element of the analysis because it determines whether conservative or non-conservative mass transfer takes place. Our results indicate that the mass of Star B is close to that of Star A and that it, too, has a very large surface He-abundance enrichment. This excludes the conservative mass transfer scenario which requires rather similar masses initially, and thus leads to large mass ratios after the Roche lobe overflow (RLO) phase (Wellstein & Langer 1999). Non-conservative interaction, on the other hand, as expected for relatively large initial mass ratios (Petrovic et al. 2005), could leave two similar mass objects after the envelope of the initially more massive star is stripped. However, the initially less massive star would still be relatively unevolved at this time, in contrast with the rather evolved state of both stars in HD 5980. Hence, the classical binary evolution scenarios cannot readily explain the parameters of HD 5980.

The question then remains whether it is possible to avoid the large radius increase in the post-MS evolutionary stages, thus avoiding the problem of mass transfer in the system, but at the same time enhancing the surface chemical composition.

Rotational mixing brings the helium and nitrogen produced in the core efficiently to the stellar surface, thereby enriching the visible surface chemical abundances without the need of stripping large amounts of mass from these outer layers. Since this rapid rotation greatly reduces the chemical gradients present in slower rotating stars, these stars undergo what is referred to as quasi-chemically homogeneous evolution (QCHE). An important characteristic of these models is that both L and T increase with time, and the star remains relatively compact, thus avoiding a RLO phase. According to Brott et al. (2011), stars with rotational velocities above $\sim$300 km s$^{-1}$ are at the threshold for QCHE at SMC metallicity.

We computed a model binary system with initial masses 90 $M_\odot$ (Primary) and 80 $M_\odot$ (Secondary), rotational velocity $v_{\text{rot}} = 500$ km s$^{-1}$, and orbital period of 12 days using the Binary Evolutionary Code (Heger et al. 2000; Petrovic et al. 2005; Yoon et al. 2006). For the early evolutionary stage, the prescription of Vink et al. (2001) for the mass-loss rate was adopted and for the WR stage we used the prescription of Hamann et al. (1995), with the correction of Vink & de Koter (2005) to account for the SMC metallicity. Because of the angular momentum loss due to the stellar winds, the orbital separation grows with time. After $\sim$3.1 Myr, the binary has an orbital period of 19.2 days and the stars have each $M \approx 67 M_\odot$, $\log(L/L_\odot) \sim 6.3$, highly enhanced surface He-abundances, and current surface temperatures of 63 kK, similar to the properties of Star A and Star B. More importantly, the evolution of these QCHE model stars places them on the HRD at systematically higher temperatures and luminosities, as shown in Figure 13, without any large radius increases that would lead to mass transfer.

The other parameters of the QCHE model stars are listed in Table 8, along with the corresponding values for Star A. Since the Secondary is the originally less massive member of the binary system, it evolves slower than the companion and

$^{15}$ See illustrations of a similar geometries in Figures 20–22 of Koenigsberger (2004).
Figure 13. Hertzsprung–Russell diagram (HRD) of our binary model with an initial orbital period of 12 days, initial surface rotational velocity $v_{\text{ini}} = 500$ km s$^{-1}$, and initial masses of 90 $M_\odot$ and 80 $M_\odot$, compared for reference to the evolution of same mass non-rotating stars. Black dots are shown every 0.1 Myr. The red circles indicate the position of the two HD 5980 binary components after $\sim$3.1 Myr, which is when their surface properties agree best with those of Star A and Star B.

(A color version of this figure is available in the online journal.)

Figure 14. Value of the electron scattering Eddington factor over time since arriving at the main sequence showing its systematic increase.

(A color version of this figure is available in the online journal.)

The computed model is with initial orbital period of 12 days, surface rotational velocity $v_{\text{ini}} = 500$ km s$^{-1}$, and metallicity $Z = 0.0021$. $v_{\text{surf}}$ is the surface rotation velocity.

Notes.

a The computed model is with initial orbital period of 12 days, surface rotational velocity $v_{\text{ini}} = 500$ km s$^{-1}$, and metallicity $Z = 0.0021$. $v_{\text{surf}}$ is the surface rotation velocity.

b References for Columns 2 and 3: (1) Mokiem et al. 2006; (2) Georgiev & Koenigsberger 2004; (3) Georgiev et al. 2011; (4) Koenigsberger et al. 1998; (5) Perrier et al. 2009.

thus corresponds to Star A, so the comparison to be made is between Star A’s observationally derived parameters and those of the secondary in the QCHE model. The mass-loss rate of Star A in 2009 is listed by Georgiev et al. (2011) as $M / \sqrt{f} = 23 \times 10^{-6} M_\odot$ yr$^{-1}$, and they adopt a wind “clumping” factor $f = 0.025$. Hence, $M_A = 3.6 \times 10^{-5} M_\odot$ yr$^{-1}$, which is very similar to that of the Secondary in the QCHE model.

Although there is no direct means of determining the surface rotation velocity of the WR-type stars in HD 5980, Georgiev & Koenigsberger (2004) speculated that the low-amplitude modulation of Star C’s photospheric absorptions that was reported in early observational data by Breysacher et al. (1982) could be caused by a much larger amplitude Doppler shift of underlying broad lines which, due to their width, would not be detectable in the HD 5980 spectrum. Under this assumption, they found that the possible absorption lines would correspond to Star A rotation speeds of $v \sin i \sim 200$–250 km s$^{-1}$, which is of the order of the rotation predicted for the Secondary in the QCHE model.

The final point to note concerns the predicted radii of the QCHE model stars, $R_{\text{prim}} = 12.5 R_\odot$ and $R_{\text{sec}} = 11.3 R_\odot$. The analysis of the optical eclipse light curve of HD 5980 (Perrier et al. 2009) yielded relative radii of $\rho_A/a = 0.158$ and $\rho_B/a = 0.108$, where $a$ is the semi-major axis of the orbit. From Table 6, $a = 151 R_\odot$ which implies that the occulting disks are $R_B \sim 16 R_\odot$ and $R_A \sim 24 R_\odot$. The latter is $\sim 2$ times larger than predicted by the QCHE model. However, it must be considered that: (1) the derived radius for the occulting disk in WR stars depends on the wavelength of observations, with smaller wavelengths yielding smaller radii (Cherepashchuk et al. 1984), so a small correction may be viable if UV light curves are analyzed; and (2) the light curve solution of Perrier et al. (2009) is based on the assumption that only Star B possesses an extended optically thick wind, so it remains to be determined whether assuming that both stars have such a region would result in a smaller radius for Star A.

We thus conclude that a plausible scenario for the evolutionary state of the 19.3 day binary in HD 5980 is one in which both stars have followed QCHE pathways.

Furthermore, the internal luminosity structure of the QCHE models is such that therein may reside the mechanism that triggered the sudden eruption observed in 1994. The increasing luminosity over time causes an increase in the Eddington electron scattering $\Gamma$-factor ($\Gamma_e$), thus lowering the effective gravity at the surface. Figure 14 shows how the $\Gamma_e$-factor grows beyond 0.5 as the age of the star approaches that of Star A. This is consistent with the value of $\Gamma_e$ found by Georgiev et al. (2011) from the spectrum of 2009. More important, however, is the fact that the total $\Gamma$ (i.e., that computed with the total opacity, not only the electron scattering opacity) becomes larger than unity near the surface layers (Figure 15). One possible mechanism driving the eruption observed in 1994 might be related to this super-Eddington layer which might turn unstable (Humphreys & Davidson 1994; Stothers & Chin 1996). We speculate that in the case of HD 5980, the instability could have been triggered by the perturbations that arise due to its eccentric orbit. A discussion of the potential role of the companion in the case of $\eta$ Car has also been discussed by Davidson (2012).

The question that then arises is whether the same mechanism could be active in other LBVs. An objection to the QCHE
scenarios is that in order for stars to follow this path, the rotation rate needs to remain high throughout their evolution, a condition that is difficult to satisfy due to the expected loss of angular momentum through stellar winds and tidal interactions. To avoid this problem, de Mink et al. (2009) suggested that very short-period, tidally locked binary systems could sustain the very rapid rotation velocities that are required. However, none of the known LBV binaries are in orbits that are close enough to meet this criterion, even after considering orbital evolution.

However, the important point to note is that for QCHE evolution to take place, the star must be able to efficiently mix the nuclear processed material up to the surface layers, and rapid rotation may not be the only mechanism that can produce this effect. For example, though speculative at this time, an alternative to rapid rotation for efficient mixing might be the time-dependent velocity field that is driven by the tidal interactions in an asynchronously rotating binary star (Koenigsberger & Moreno 2013). With this mechanism, one could start out with a relatively wider binary than is required by the fast-rotation models and, as long as the stellar rotation remains asynchronous, the induced velocity field could contribute toward the mixing processes. Whether the efficiency with which the mixing occurs is sufficient for QCHE to be attained, however, remains to be proven.

Little can be said at this time about the evolutionary status of the Star C binary except for noting that such highly eccentric binary systems are rarely found in orbits with periods shorter than 100 days, which suggests that it may be gravitationally bound to the Star A + Star B pair. The 5:1 resonance in the orbital periods of the two systems is further evidence in support of this hypothesis. In triple hierarchical systems, the outer period tends to be a multiple of the inner binary period, as in $\alpha$ Ori (Aa+Ab) + B (Simón-Díaz et al. 2011). If a similar condition were to hold in HD 5980, then both $P_{AB}$ and $P_c$ would be sub-multiples of the outer period and hence, would be related to each other, as observed. Thus, we speculate that the Star C components have a similar age as Star A and Star B.

7. CONCLUSIONS

Due to their short lifetime, LBVs are quite rare and most of them are apparently single objects, impeding a direct determination of their current mass. Determining basic parameters for the LBVs undergoing major eruptions is even more challenging, since the majority are detected only in distant galaxies and little or no information exists on the properties of the star prior to the eruptive event.

In this context, the importance of HD 5980 for understanding massive star evolution cannot be overstated. It provides the closest and clearest example of mass-loss phenomena during the short-lived LBV and WR phases in low-metallicity star forming regions such as those where the most distant LBV events are being observed and which, in addition, host the brightest gamma-ray events. Furthermore, massive objects such as those in HD 5980 are believed to be the progenitors of some of these events.

The following are the conclusions of this paper.

1. The masses in the 19.3 day Star A+Star B system have been constrained under the premise that both stars currently possess a WNE-type spectrum. The largest uncertainties in this determination derive from: a) the fact that the spectrum of Star B has not been unambiguously isolated from that of Star A; and b) the lack of detailed models predicting the strength of the emission lines that can arise in the WWC region.

2. A plausible scenario for the evolution of the 19.3 day binary in HD 5980 is one in which both stars have followed a quasi-chemically homogeneous evolutionary path. We speculate that the high mixing efficiency that is required may be provided, at least in part, by the non-uniform differential rotation structure that is driven by the tidal forces in an asynchronously rotating binary star. HD 5980 may thus be the first binary system which provides evidence that QCHE can indeed occur, since it possesses the two extreme properties that are required for this scenario, i.e., a very high mass of its components, and a very low metallicity. If confirmed this would imply that both stars in HD 5980 can be considered as candidates for gamma-ray burst progenitors (Yoon et al. 2006) or even pair instability SNe (Langer et al. 2007).

3. The instability that caused the 1994 eruption may be related to the super-Eddington layer which is present just below the photosphere in the QCHE models.

4. A 96.5 day period is confirmed for the RV variations of the photospheric absorptions arising in the third component, Star C, indicating that this star is also a binary. We point out that this period is in a 5:1 ratio with the period of the Star A + Star B binary, suggesting that these four stars may constitute a hierarchical quadruple system.

We dedicate this paper to the memory of our friends and colleagues Virpi Niemela and Leonid Georgiev.

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