A stripped-companion origin for Be stars: clues from the putative black holes HR 6819 and LB-1

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1 INTRODUCTION

Large-scale multi-epoch radial velocity (RV) surveys have identified a number of binaries in recent years that are proposed to contain stellar mass black holes (BHs; e.g. Casares et al. 2014; Khokhlov et al. 2018; Giesers et al. 2018, 2019; Thompson et al. 2019). Confirmation of these BH candidates is challenging precisely because BHs are expected to be rare. That is, all plausible alternate explanations for observed data, even those which are rare, must be ruled out before a candidate BH can be considered reliable.

Recently, Rivinius et al. (2020) identified the object HR 6819 as a candidate host of a stellar-mass BH. Phase-resolved optical spectra of the object revealed two luminous components: a B star with relatively narrow, RV-variable absorption lines, which were observed to follow a nearly circular orbit with $P = 40.3$ days and velocity semi-amplitude $K_B \approx 61 \text{ km s}^{-1}$, and a classical Be star whose broad emission and absorption lines appeared to be stationary. No component was found to orbit in anti-phase with the B star.

Assuming the mass of the B star to be at least $5 M_\odot$, as expected for a normal star of its spectral type, Rivinius et al. (2020) argued that any stellar companion massive enough to explain the B star’s orbit would also contribute to the spectrum at a detectable level. They thus concluded that the companion is a BH, with an estimated minimum mass of $4.2 M_\odot$. In this hierarchical triple scenario, the Be star must be at least a few AU from the B star and BH for the system to be dynamically stable, and its status as a Be star would likely be unrelated to the B star or BH.

HR 6819 is in many ways similar to the binary LB-1 (Liu et al. 2019), which was also proposed to contain a stellar-mass BH. Like HR 6819, LB-1 contains an RV-variable B star and apparently stationary emission lines. The emission lines were initially proposed to originate in an accretion disk, either around the BH (Liu et al. 2019) or around the binary (El-Badry & Quataert 2020; Abdul-Masih et al. 2020; Irgang et al. 2020). Rivinius et al. (2020) proposed that LB-1 and HR 6819 are both hierarchical triples with a B star and a BH companion in the inner binary, and a distant Be star – the source of the emission lines – orbiting both components.

Shenar et al. (2020), however, recently used spectral disentangling to fit the multi-epoch spectra of LB-1 as a sum of two luminous components. They also found evidence for a Be star, including both
emission and rotationally-broadened absorption lines, but found its spectrum to shift in anti-phase with the B star. This is not expected in the hierarchical triple scenario, so they argued that LB-1 is a binary containing two luminous stars. It was also recently noted by Liu et al. (2020) that the emission line shape in LB-1 varies coherently with the B star’s phase. This variation, which is consistent with expectations for an irradiated or tidally-perturbed disk, is expected if the system is a binary, but not if it is a hierarchical triple.

The RV variability amplitude found by Shenar et al. (2020) for the Be star in LB-1 suggests it is 5 times more massive than the B star. If the luminous binary scenario for LB-1 is correct, the B star’s phase. This variation, which is consistent with expectations for an irradiated or tidally-perturbed disk, is expected if the system is a binary, but not if it is a hierarchical triple.

In this paper, we use spectral disentangling to fit the multi-epoch spectra of HR 6819. We find that, like in LB-1, the Be star orbits in anti-phase with the B star, which is much less massive than expected for a normal star of its spectral type. The remainder of this paper is organized as follows. We describe the spectra in Section 2.1 and the disentangling method in Section 2.2. We constrain the atmospheric parameters of both components in Section 2.3. Section 2.4 presents the RV variability of both components, and Section 2.5 compares the system’s luminosity to models. The emission from the Be star’s disk is examined in Section 2.6, and the abundances of the B star in Section 2.7. We present possible evolutionary models for the system in Section 2.8. Finally, we discuss HR 6819 and LB-1 in the context of the broader Be star population in Section 3.

Appendix A describes our estimate of both components’ masses, radii, and luminosities. Spectral disentangling is detailed in Appendix B. Appendix C investigates the pulsation-driven variability of the B star. Spectroscopic constraints on the B star’s rotation velocity are presented in Appendix D.

## 2 METHODS

### 2.1 Data

We analyze 51 optical spectra of HR 6819 with spectral resolution \( R \approx 48,000 \) that were obtained in 2004 with the FEROS echelle spectrograph (Kaufer et al. 1999) on the ESO/MPG 2.2m telescope at La Silla Observatory. The data span 134 days (3.3 orbital periods) from MDJ 53138 to 53273 and are publicly available through the ESO archive; they are described in more detail by Rivinius et al. (2020). The spectra were reduced with the ESO-MIDAS pipeline, which performs bias-subtraction, flat fielding, and wavelength calibration, applies a heliocentric correction, and combines the spectra from individual orders. The pipeline failed to reliably measure the signal-to-noise ratio (SNR), so we estimated it empirically from the pixel-to-pixel scatter in regions without strong absorption or emission lines. The typical SNR per pixel is 300 at \( \lambda = 6,000 \text{ Å} \). We verified the stability of the wavelength solution by checking that the ISM sodium absorption line at 5890 Å is found at the same heliocenter-corrected wavelength at all epochs.

We also analyzed an additional 12 FEROS spectra of HR 6819 that were obtained in 1999 (dataset “A” in Rivinius et al. 2020). We found that the Be star’s emission line profiles varied substantially between the 1999 and 2004 datasets, with obvious changes in the shape, width, and amplitude of Balmer, Fe II, and O I emission lines. Such variation is common in Be stars due to changes in the structure of the disk (e.g. Dachs et al. 1981; Okazaki 1991), but it complicates the disentangling of the composite spectra. We therefore focus our analysis on the 2004 dataset, within which the emission line profiles are relatively stable.

The spectra are not flux calibrated. Continuum normalization was performed by fitting a cubic spline to wavelength pixels without significant absorption or emission. The initial set of candidate continuum pixel was selected by taking all wavelengths from a TLUSTY model spectrum (Hubeny & Lanz 1995) with \( Teff = 18 \text{ kK} \) and \( \log g = 3.5 \) for which all pixels within \( \pm 600 \text{ km} \text{s}^{-1} \) are within 0.5% of the theoretical continuum. We then refined the selection of continuum pixels by inspecting the individual spectra, removing pixels affected by telluric absorption and other features not present in the model spectrum. We tested application of this continuum normalization procedure on mock spectra with a range of \( \log g \) to verify that it does not significantly overfit the continuum in the wings of broad lines. This is important, because improper continuum normalization can change the shape of the wings of broad lines, leading to biases in the inferred \( \log g \). We focus our analysis on the part of the spectrum with \( \lambda > 3900 \text{ Å} \) because the narrow spacing of the Balmer lines at shorter wavelengths makes it challenging to estimate the continuum there.

### 2.2 Spectral disentangling

We used the code CRES (Ilijic 2004), which implements the wavelength-space spectral disentangling algorithm of Simon & Sturm (1994), to separate the spectra of the two components. The algorithm operates under the ansatz that all the observed composite spectra are produced by summing together the time-invariant spectra of two components, with known Doppler shifts between epochs and a known continuum luminosity ratio. The rest-frame spectra of the two components are modeled as arbitrary vectors, which under these assumptions can be solved for using singular value decomposition. No model spectra (or indeed, any physics beyond the Doppler shift) are used in this calculation.

The velocity of the B star in HR 6819 is reasonably well known at each epoch through the orbital solution from Rivinius et al. (2020), but the velocity of the Be star is not. Following Shenar et al. (2020), we therefore proceed under the ansatz that the Be star and B star are orbiting each other, and step through a grid of velocity semi-amplitudes for the Be star, \( Ke_B \), calculating the best-fit disentangled spectra for each value. The choice of \( Ke_B = 0 \text{ km} \text{s}^{-1} \) corresponds to a stationary Be star and would be expected in a triple scenario. We choose the optimal \( Ke_B \) and spectral decomposition as the value that minimizes the total \( \chi^2 \) over all spectra; we find \( Ke_B = (4.5 \pm 2) \text{ km} \text{s}^{-1} \). The determination of this value is described in detail in Appendix B.

When determining \( Ke_B \), we use the spectra at their native resolution and independently analyze narrow wavelength ranges centered on absorption and emission lines with clear contributions from both components. We then adopt \( Ke_B = 4.5 \text{ km} \text{s}^{-1} \) and use this value to disentangle the full spectrum. At this stage, we re-bin the spectra to a 0.2 \( \text{Å} \) wavelength resolution to reduce the computational cost of disentangling. This does not significantly decrease the information content of the spectra, because higher-frequency spectral features are smeared out by rotation and/or macroturbulence. The spectral disentangling and determination of \( Ke_B \) are described in detail in Appendix B.
**HR 6819: A stripped star + Be star binary**

Spectral decomposition only determines the spectra of the two components up to a multiplicative constant, which represents the continuum luminosity ratio of the two stars. Unless there are e.g. eclipses or well-characterized pulsations, the luminosity ratio must be determined based on external information, because the effects on the combined spectrum of increasing the continuum flux contribution of one component are identical to the effects of weakening all its spectral lines. We begin by assuming both components contribute 50% of the light, as estimated by Rivinius et al. (2020). Once a preliminary decomposition is found with this ratio, we adjust the luminosity ratio to match the depths of the B star’s Balmer lines with the value theoretically predicted based on the effective temperature measured from Si and Fe ionization balance (see Section 2.3; this temperature estimate is not sensitive to the luminosity ratio). We found the optimal flux ratio to be $f_B/f_{tot} = 0.47 \pm 0.05$ at $\lambda = 4000 \AA$ (i.e., with the Be star the slightly brighter component). Because the Be star is slightly hotter, its relative contribution to the total flux is expected to decrease toward redder wavelengths. For our fiducial temperature estimates of $T_{\text{eff}, \text{B}} = 16$ kK and $T_{\text{eff}, \text{Be}} = 18$ kK (Section 2.3), the expected flux ratio increases to 0.49/0.51 at $\lambda = 7000 \AA$, still with the Be star the brighter component. This weak wavelength-dependence of the flux ratio is accounted for in our spectral disentangling. We normalize the component following the method described by Ilijic et al. (2004).

Figure 1 shows a 600 Å wide section of the disentangled spectra. The overplotted TLUSTY model spectra are taken from the BSTAR06 grid (Lanz & Hubeny 2007) and have the $T_{\text{eff}}$, log $g$, and $v \sin i$ that we find to best fit the disentangled spectra. We note that these assume the Solar abundance pattern; we show in Section 2.7 that the surface abundances of the B star are somewhat unusual. The Be star spectrum primarily shows strong absorption lines from neutral H and He due to its large $v \sin i$, as well as excess due to double-peaked emission lines within the Balmer line cores and in Fe II lines (Section 2.6). The B star spectrum has many narrow metal lines (see Section 2.7).

Fitting the disentangled metal line profiles of both components with the rotational profile from Gray (1992), we find $v \sin i = (180 \pm 20) \text{km s}^{-1}$ for the Be star. For our favored Be star mass, radius, and inclination (Sections 2.4 and 2.5), this corresponds to near-critical rotation. For the B star, we derive an upper limit of $v \sin i < 20 \text{km s}^{-1}$. We are not able to measure the actual value of $v \sin i$ from the B star’s line profiles because non-rotational broadening due to macroturbulence and/or non-radial pulsations dominates over rotation (see Appendix D). If the B star is tidally synchronized, it might be expected following a period of mass transfer, its $v \sin i$ would be less than 5 km s$^{-1}$.

Overall, the consistency between the disentangled spectra and theoretical models, as manifest in properties such as the relative depth and width of absorption lines, is good. We emphasize that the spectral disentangling procedure does not use model spectra in any way – the disentangled spectra are free to take any shape they like – so the agreement between the observed and model spectra is encouraging.

### 2.3 Temperature and gravity

We fit the disentangled spectra of both components using 1D-NLTE model spectra computed with TLUSTY (Hubeny & Lanz 1995, 2017). We use the BSTAR06 grid, which has $\Delta T_{\text{eff}} = 1$ kK and $\Delta \log g = 0.25$ dex, (Lanz & Hubeny 2007) in an initial grid search and then use the radiative transfer code SYNSPEC (Hubeny & Lanz 2011) to generate spectra with finer grid spacing, as well as spectra with different abundance patterns than assumed in the BSTAR06 grid (Section 2.7). We perform an iterative fit for the temperature and gravity of the B star, first measuring $T_{\text{eff}}$ based on equivalent width ratios of lines in different ionization states (Figure 3), then measuring log $g$ based on the wings of broad lines, then refining the $T_{\text{eff}}$ estimate taking the estimated log $g$ into account, and so forth. Detailed abundances are measured in a final step, once a converged solution for $T_{\text{eff}}$ and log $g$ is found. To estimate systematics due to model uncertainties, we also compared to 1D-LTE model spectra generated with SYNTHE (Kurucz 1993) from ATLAS-12 model atmospheres (Kurucz 1970, 1979, 1992). We find consistent solutions for the two sets of model spectra within 1 kK in $T_{\text{eff}}$ and 0.1 dex in log $g$.

Our final best-fit parameters for both stars are listed in Table 1. A critical result of our analysis is that the surface gravity of the B star is lower than that of a main-sequence star, log $g \approx 2.75$. The main constraint on log $g$ comes from the wings of the Balmer lines and some He I lines, which become broader with increasing log $g$. This is illustrated in Figure 2, which shows a zoom-in on the Balmer lines He I, Hδ, and Hγ in the disentangled spectrum of the B star. We compare the best-fit model, which has log $g = 2.75$, to one with log $g = 3.75$, the approximate value assumed by Rivinius et al. 2020, and the best-fit $T_{\text{eff}}$ for that log $g$. The log $g = 3.75$ model predicts much wider Balmer lines than are found in the observed spectrum. The poor fit of the higher log $g$ model is also evident in the wings of the 4009 Å and 4388 Å He I lines. We note that while the superiority of the log $g = 2.75$ model is clear in Figure 2, this is only the case because we have disentangled the spectra of the two components. In the composite spectra, the Balmer and He I lines appear much broader due to the broader absorption lines of the Be star.

We measure the effective temperature of the B star from equivalent width ratios of lines of the same element in different ionization states. Higher ionization states become increasing populated as temperature increases. The ratios of e.g. Si III to Si II line strengths are therefore sensitive to $T_{\text{eff}}$ and insensitive to the absolute Si abundance, since changing the abundance changes the number of atoms in both ionization states. Unlike absolute abundance measurements, equivalent width ratios are also insensitive to the assumed luminosity ratio of the two components, since light from the companion dilutes all lines by a similar factor.

Figure 3 compares the observed equivalent width ratios of Si III/II and Fe III/II lines to predictions from TLUSTY models for a range of $T_{\text{eff}}$ and log $g$. Our preferred model with log $g = 2.75$ is shown in black; colored lines show other values of log $g$. All the line ratios we consider imply a temperature of $T_{\text{eff}} = (16 \pm 0.5)$ K K for log $g = 2.75$, with values that are respectively higher and lower for higher or lower log $g$.

Rapid rotation washes out most of the Be star’s lines, making it infeasible to measure $T_{\text{eff}}$ from ionization state ratios. The constraint on $T_{\text{eff}}$ comes primarily from the Balmer lines and the strength of Si II and He I lines. The star’s gravity can still be constrained from the shape of the Balmer lines, which are sufficiently broad that their wings are not strongly affected by rotation.

A natural concern when working with echelle spectra is that the profiles of broad lines such as the Balmer lines could be distorted by the blaze function and/or continuum normalization process, leading to a bias in log $g$. To assess the reliability of our derived $T_{\text{eff}}$ and log $g$, we analyzed an archival FEROS spectrum of the slowly rotating standard B star HR 5285 that was presented and analyzed by Nieva & Przybilla (2007). The spectral resolution and SNR of this spectrum is similar to that of the HR 6819 data.
et al. (2020). RVs measured from emission lines can be biased if lines, which likely originate in a disk around the Be star. Absorption lines from the Be star but contains seven Fe II emission components left free. We also leave the luminosity ratio free, because at least one component of the system is photometrically variable at the 5-10% level (Rivinius et al. 2020). We mask regions of the spectrum containing diffuse interstellar bands, telluric absorption, and a few lines that are poorly fit.

We divide the spectrum into four segments, fitting the RVs in each segment independently. This allows us to empirically estimate RV uncertainties from the scatter between different segments. It also allows us to test whether the RVs of the emission lines, which trace the Be star’s disk and circumstellar envelope, vary coherently with the absorption lines, which trace the photosphere. In particular, the region of the spectrum with 5100 < λ/Å < 5400 contains no strong absorption lines from the Be star but contains seven Fe II emission lines, which likely originate in a disk around the Be star.

As discussed by El-Badry & Quataert (2020) and Abdul-Masih et al. (2020), RVs measured from emission lines can be biased if contamination from absorption lines is not taken into account. Our procedure for measuring RVs is not susceptible to the bias pointed out in these works, because the spectra of both components are modeled simultaneously.

Figure 4 shows the measured epoch RVs for both components, phased to a period of 40.3 days. Error bars in the top two panels show the scatter between RVs fit to different spectral regions. The Be star displays sinusoidal RV variability with the same period as the B star, suggesting that they are orbiting each other. Emission and absorption lines show consistent variability, implying that the Be star and its disk move coherently. This implies that the disk is likely circumstellar, not circumbinary. For the B star, the RV uncertainties are smaller than the symbols for most points, suggesting that there is intrinsic RV scatter.

We fit the RVs of both components independently with Keplerian orbits, without the requirement that they have the same center-of-mass velocity or phase. Along with the 6 standard Keplerian orbital parameters, we fit for an intrinsic RV scatter or “jitter” term (see El-Badry et al. 2018, their Equation 9), which can represent either intrinsic RV variability not captured in the model or underestimated uncertainties. We initially search for the best-fit orbital solution using simulated annealing, and then sample from the posterior in the vicinity of this solution using a Markov chain Monte Carlo method, as described in El-Badry et al. (2018). Given the large number of RV measurements (51), there is little danger of the sampler getting stuck in a local minimum. Our best-fit orbital parameters are listed in Table 1.

Our best-fit velocity amplitude for the B star is $K_B = 62.7 \pm 1.0 \text{ km s}^{-1}$, which is consistent with the value measured by Rivinius et al. (2020). Our constraints on $P$ is also consistent, but our uncertainties are larger, because we did not include the additional RVs.

Figure 1. Disentangled spectra of the B star (top) and Be star (bottom) in HR 6819. Black lines show the reconstructed spectra. Red lines show TLUSTY model spectra, which are not used in the spectral disentangling, for comparison. The B star spectrum contains many narrow lines and is well-fit by a model with $T_{\text{eff}} = 16\, \text{kK}$ and $\log g = 2.75$ (see also Figure 2). The Be star spectrum is strongly broadened by rotation. It is reasonably well described by a model with $T_{\text{eff}} = 18\, \text{kK}$ and $\log g = 3.75$, but with clear excess emission inside the Balmer line cores (see also Figure 6).
from 1999 in our fit. Unlike Rivinius et al. (2020), we do not measure any significant eccentricity: we find an upper limit of $e < 0.03$ for the B star, but the orbit is consistent with $e = 0$ once RV scatter is accounted for. For the Be star, we measure $e < 0.13$, again consistent with a circular orbit. The formal uncertainty in $K_{Be}$ obtained from fitting the RVs shown in Figure 4 is only $0.3 \text{ km s}^{-1}$, but the true uncertainty is larger, since the points shown in Figure 4 were obtained after fixing $K_{Be} = 4.5 \text{ km s}^{-1}$ in spectral disentangling (see Appendix B).

For the B star, we find an intrinsic RV scatter of $\approx 4 \text{ km s}^{-1}$. We did not find evidence of additional periodicity in the residuals. As we show in Appendix C, the B star’s RV scatter is accompanied by modest variability in its spectral line profiles. HR 6819 also exhibits photometric variability with brightness variations of up to five percent in the total light (implying variations of $10\%$ if the light comes primarily from one component) on timescales down to half a day (Rivinius et al. 2020). The most likely explanation is that the B star is a slowly pulsating B star (SPB star; e.g. Waelkens 1991) undergoing non-radial g-mode pulsations that cause the RV scatter, line profile variability, and photometric variability. The B star’s atmospheric parameters, the amplitude of the RV and photometric variability, and the variability timescale are all within the normal range found for SPB stars (e.g. Dziembowski et al. 1993).

We note, however, that photometric variability is also common for Be stars, due to both pulsations and variations in the structure of the disk (Rivinius et al. 2013), so some of the observed photometric variability could also come from the Be star.

If the masses of both components are known, the inclination of the binary’s orbital plane can be measured from the velocity

| Table 1. Physical parameters and $1\sigma$ (middle 68%) uncertainties for both components of HR 6819. Constraints on stellar parameters are based on the measured $T_{\text{eff}}$ and $\log g$ of the Be star and the dynamical mass ratio. Table A1 lists constraints that also take the distance into account. |
|-----------------|-----------------|
| **Parameters of the B star** | **Parameters of the Be star** |
| $T_{\text{eff}}$ [KK] | $T_{\text{eff}}$ [KK] |
| $\log g$ [cm s$^{-2}$] | $\log g$ [cm s$^{-2}$] |
| $\nu$ sin $i$ [km s$^{-1}$] | $\nu$ sin $i$ [km s$^{-1}$] |
| $v_{\text{mac}}$ [km s$^{-1}$] | $v_{\text{mac}}$ [km s$^{-1}$] |
| $M$ [M$_{\odot}$] | $M$ [M$_{\odot}$] |
| $R$ [R$_{\odot}$] | $R$ [R$_{\odot}$] |
| $\log(L/L_{\odot})$ | $\log(L/L_{\odot})$ |
| $\nu_{\text{crit}}/\nu_{\text{rot}}$ | $\nu_{\text{crit}}/\nu_{\text{rot}}$ |

| **Parameters of the binary** | **Be star disk parameters** |
|----------------------------|----------------------------|
| Orbital period [day] | $P$ [day] |
| $K_{\text{B}}$ [km s$^{-1}$] | $P$ [day] |
| $K_{\text{Be}}$ [km s$^{-1}$] | $P$ [day] |
| $\gamma_{\text{Be}}$ [km s$^{-1}$] | $P$ [day] |
| $\gamma_{\text{Be}}$ [km s$^{-1}$] | $P$ [day] |
| $\varepsilon_{\text{Be}}$ | $\varepsilon_{\text{Be}}$ |
| $s_{\text{B}}$ [km s$^{-1}$] | $s_{\text{B}}$ [km s$^{-1}$] |
| $s_{\text{Be}}$ [km s$^{-1}$] | $s_{\text{Be}}$ [km s$^{-1}$] |
| $q = M_{\text{Be}}/M_{\text{B}}$ | $q = M_{\text{Be}}/M_{\text{B}}$ |
| $\alpha$ | $\alpha$ |
| $\alpha$ | $\alpha$ |

| **B star abundances relative to Solar** | **Be star disk parameters** |
|---------------------------------------|----------------------------|
| [He/H] | $v_{\text{out}}$ sin $i$ [km s$^{-1}$] |
| [C/H] | $R_{\text{inner}}/R_{\text{outer}}$ |
| [N/H] | $R_{\text{inner}}/R_{\text{outer}}$ |
| [O/H] | $R_{\text{inner}}/R_{\text{outer}}$ |
| [Fe/H] | $R_{\text{inner}}/R_{\text{outer}}$ |

**Figure 2.** Balmer lines as a surface gravity diagnostic. Black lines show the spectrum of the B star obtained through spectral disentangling. Red lines shows a TLUSTY model spectrum with $\log g = 2.75$, which reproduces the observed Balmer line profiles reasonably well. Cyan lines show model spectra with $\log g = 3.75$, as expected for a normal B star near the end of the main sequence. For this value of $\log g$, the wings of the Balmer and He I lines are much too wide to match the observed spectrum. The model $T_{\text{eff}}$ values are chosen to match the observed ionization ratios at a given $\log g$ (Figure 3). For other choices of $T_{\text{eff}}$, the disagreement between the $\log g = 3.75$ model and the observed spectrum is even more severe.
Figure 3. Temperature of the B star from ionization equilibrium. Shaded regions show the observed ratios of the equivalent widths of Si III and Si II lines (left columns) and Fe III and Fe II lines (right columns). Lines show the ratios predicted for models of different log g as a function of T\text{eff}. Given the best-fit log g ≈ 2.75 from the Balmer lines (Figure 2), the best-fit T\text{eff} lies between 15500 and 16500K for all line ratios.

semi-amplitudes. For a circular orbit,

\[
\sin^3 i = \frac{(1 + K_{\text{Be}}/K_B)^2 K_B^3 P}{2\pi G M_{\text{Be}}}.
\]  

(1)

Taking \( M_{\text{Be}} = 6.7^{+1.9}_{-1.5} \) \( M_\odot \), a constraint which we obtain by comparing the Be star’s temperature and surface gravity to MIST evolutionary tracks (see Appendix A; Dotter 2016; Choi et al. 2016), we obtain \( i = 32.1^{+3.0}_{-2.6} \) deg. The mass ratio of the two components is constrained from the relative velocity amplitudes to \( M_B/M_{\text{Be}} = 0.071 \pm 0.032 \). This leads to \( M_B = 0.47^{+0.28}_{-0.22} M_\odot \), where the dynamical mass ratio and the mass of the Be star contribute roughly equally to the uncertainty.

2.5 Luminosity

If the distance to HR 6819 is known, then its absolute magnitude can distinguish between models in which the B star is a normal star and models in which it was recently stripped. The bolometric
Figure 4. Radial velocities for the B star (top panel) and Be star (middle and bottom panels), phased to a period of 40.3 days. RVs for both components are measured simultaneously by fitting the composite spectra with the single-star templates obtained from spectral disentangling. The Be star exhibits sinusoidal RV variability in anti-phase with the B star, suggesting the stellar photosphere) and emission lines (presumably originating in a circumstellar disk). Emission and absorption line RVs vary coherently.

As discussed by Rivinius et al. (2020), a normal star with the temperature and gravity of the B star in HR 6819 would have a minimum mass of at least $5 \, M_\odot$. It would thus be a minimum of 10 times brighter than a $0.5 M_\odot$ stripped star with the same $T_{\text{eff}}$ and $\log g$. And as we discuss below, the minimum mass for a normal B star matching HR 6918’s spectral type is even higher, of order $10 \, M_\odot$, once the lower log $g$ measured from the disentangled spectrum is taken into account.

2.5.1 Distance

Gaia measured a parallax of $\pi = 2.91 \pm 0.18 \, \text{mas}$ for HR 6819 (Gaia DR2 source id 6649357561810851328), implying a distance between 324 and 366 pc (Gaia Collaboration et al. 2018). At this distance, the projected maximum separation of the two components would be $\theta \approx 1.25 \, \text{mas}$. Since this is almost half the parallax and the astrometric solution did not account for orbital motions, the parallax uncertainty is likely underestimated. Because the two components of HR 6819 have almost the same brightness, the photocenter wobble is expected to be somewhat less than the angular size of the orbit. Given that the Gaia data span many orbital periods, the Gaia distance is also not expected to be catastrophically wrong.

Rivinius et al. (2020) derived an independent estimate of the distance to HR 6819 by comparing its flux-calibrated UV spectrum to models with temperature and radius matching their fit to the spectrum. They estimated $d = (310 \pm 60)$ pc, which is also consistent with the Gaia distance. The largest source of uncertainty was the unknown luminosity ratio between the two components. We adopt this same estimate in our analysis. Although the spectral type we derive for the B star is significantly different from that estimated by Rivinius et al. (2020), only the temperature and radius are relevant for the distance estimate. The values they adopted ($T_{\text{eff}} = 16 \, \text{kK}$ and $R = (5.5 \pm 0.5) R_\odot$) are consistent with our best-fit values, because our lower mass estimate is almost exactly compensated for by our lower estimate of $\log g$.

2.5.2 Synthetic photometry

We used synthetic photometry from MIST isochrones (Dotter 2016; Choi et al. 2016) to place models for HR 6819 on the color-magnitude diagram (CMD), assuming solar metallicity. We assume an extinction of $E(B-V) = 0.135$, as reported by Wegner (2002), and a Cardelli et al. (1989) extinction curve with $R_V = 3.1$. The SFD map estimates $E(B-V) = 0.13$ at the position HR 6819 (Schlegel et al. 1998).

We began by searching the MIST model grid for “normal” stellar models with atmospheric parameters similar to the B star ($15.5 < T_{\text{eff}}/\text{kK} < 16.5$ and $2.7 < \log g < 3$). These models have masses between 9 and $14 \, M_\odot$ (see Figure A1). This is significantly higher than the mass of $(5-6) \, M_\odot$ that would be inferred if the B star had $\log g = 3.5-4$, as assumed by Rivinius et al. 2020. These models are shown in green in the right panel of Figure 5. Even without the contribution from the Be star, these models are far brighter than HR 6819.

We next calculate synthetic photometry for a stripped star with the same range of $T_{\text{eff}}$ and $\log g$ as the models described above, accounting for the lower mass using Equation 3. These models are shown in magenta in Figure 5; we assume masses between 0.4 and $0.6 \, M_\odot$ for the stripped star. We also take synthetic photometry for the Be star from models with $15 < T_{\text{eff}}/\text{kK} < 20$ and $3.75 < \log g < 4.25$; these models are shown in cyan. Finally, we calculate the combined magnitude and color of randomly paired Be star and stripped star models (orange points). These are generally consistent with the color and magnitude of HR 6819.

The poor agreement between a “normal B star” model and the observed luminosity of HR 6819 cannot be alleviated by simply assuming a different distance. A distance of 3-5 times our fiducial value would make the observed magnitude of HR 6819 compatible with a normal B star (green points in Figure 5). But in this case, the B star would be a factor of $\sim 20$ brighter than the Be star, violating
the observational constraint that the two components are roughly equal.

2.5.3 Luminosities, masses and radii

We estimate the mass and radius of the Be star by comparing its temperature and surface gravity to a grid of MIST evolutionary tracks (Figure A1). We then calculate the mass of the B star from the measured mass ratio, and its radius from this mass and its measured surface gravity. The bolometric luminosity of each component is then calculated through Equation 2. The resulting stellar parameters are reported in Table 1.

In Appendix A, we calculate an estimate of the same stellar parameters that is anchored on the distance to HR 6819 rather than on evolutionary tracks. We obtain consistent solutions between the two methods.

2.6 Emission lines

As is often the case in classical Be stars, the disentangled Be star spectrum contains clearly double-peak emission in the Balmer series. Weak double-peak emission lines are also found in some Fe II and Si II lines. Figure 6 highlights the Hα line (which is by far the strongest of the star’s emission lines) and several of the Fe II lines. We correct the emission lines for contamination from the Be star’s absorption lines by subtracting a TLUSTY model spectrum with $T_{\text{eff}} = 16$ K and $\log g = 2.75$, varying each of the 11 elements listed in Table 1 with a grid spacing of 0.1 dex. All abundances are measured relative to the Solar abundance pattern, which is taken from Grevesse & Sauval (1998). The atmospheric structure is taken from the BSTAR06 atmosphere grid, which was generated with TLUSTY assuming $Z = 0.02$ and the Solar abundance pattern, with a microturbulent velocity of $2 \text{ km s}^{-1}$.

Unlike the line ratios shown in Figure 3, the abundances are sensitive to the luminosity ratio of the two components. The smaller the ratio $L_B/L_{\text{Be}}$, the deeper the lines of the B star, and the larger the inferred abundances. Our reported abundance uncertainties account for a ±5% uncertainty in the luminosity ratio; this dominates the uncertainty for most elements. We also note that the inferred abundances are moderately sensitive to the assumed microturbulent velocity, but we do not vary it in our analysis.

The most striking result of our abundance analysis is that the surface of the B star is significantly enriched in nitrogen and helium, and significantly depleted of carbon. This suggests that material on the B star’s surface has been processed by the CNO cycle, as has also been reported for LB-1 (Irrgang et al. 2020; Shenar et al. 2020). An enhancement of nitrogen and helium at the expense of carbon and oxygen is expected in the stripped star scenario: the material at the surface of the star today was inside the convective core of the progenitor when it was on the main sequence and is thus contaminated with fusion products.

Unlike in LB-1, we do not find a significant depletion of oxygen for the B star. Most other elements are consistent with, or slightly enhanced relative to, the Solar value. We note that Irrgang et al. (2020) found most metals to be depleted in LB-1, but this is likely at least in part due to the unaccounted-for line dilution by the Be star, which makes all lines appear weaker by roughly a factor of 2.

We do not attempt to measure detailed abundance for the Be star, since most metal lines are washed out by rotation, uncertainties in continuum normalization will significantly affect the interpretation of weak and highly broadened lines, and contamination by disk emission complicates the interpretation of absorption line equivalent widths. Besides H and He, the strongest lines visible in the Be star spectrum come from C II and Si II. The strongest tension occurs for the C II line, which is much weaker than predicted. This suggests that the Be star’s surface is also depleted of C, which is consistent with our expectations if the Be star recently accreted much of the CNO-processed envelope of the B star.

We defer more detailed analysis and modeling of HR 6819’s abundance pattern to future work. Given the B star’s unusually large surface He abundance, it would be useful to recalculate the atmospheric structure while taking the abundance variations into account.

2.8 Evolutionary history

To better understand the possible formation pathways and future evolution of HR 6819, we began by searching the BPASS (v2.2;
Eldridge et al. (2017) library of binary evolution models for models that go through a phase similar to the current state of HR 6819. Because BPASS models do not follow the evolution of the secondary with detailed calculations, we then investigated a few possible formation histories in more detail using Modules for Experiments in Stellar Astrophysics (MESA, version 12778; Paxton et al. 2011, 2013, 2015, 2018, 2019).

We search the $Z = 0.02$ BPASS grid and set the following constraints:

- $15 < T_{\text{eff},1}/kK < 17$
- $2.5 < \log g_1 < 3.25$
- $15 < T_{\text{eff},2}/kK < 20$
- $5 < M_2/M_\odot < 8$
- $20 < P/\text{days} < 60$
- $2.7 < \log(L_1/L_\odot) < 3.3$

This search yielded 10 models, with initial primary masses ranging from $4.5$ to $7 M_\odot$, initial mass ratios ranging from $0.4$ to $0.8$, and initial periods of $1.5$ to $6.5$ days. We then used MESA to calculate binary evolution tracks for models in this region of parameter space, and to investigate how the systems’ evolution changes when their parameters are varied. Most of the single-star free parameters are set following the inlist_7M_prem_7to_AGB.inlist in the MESA test suite. We use the wind mass loss prescription from Reimers (1975). The initial metallicity is set to 0.02. We use a mixing length parameter $\alpha_{\text{MLT}} = 1.73$ and an exponential overshoot scheme with overshoot efficiency parameter $f_{\text{ov}} = 0.014$. The overshooting scheme and motivation for this choice of $f_{\text{ov}}$ are described in Herwig (2000).

The MESA binary module is described in detail in Paxton et al. (2015). Both stars are evolved simultaneously, with tides and mass transfer taken into account. Roche lobe radii are computed using the fit of Eggleton (1983). Mass transfer rates in Roche lobe overflowing systems are determined following the prescription of Kolb & Ritter (1990). Tidal torques are modeled following Hut (1981). The orbital separation evolves such that the total angular momentum is conserved when mass is lost through winds or transferred to a companion, as described in Paxton et al. (2015). We experimented with a range of mass transfer efficiencies. When mass transfer is not conservative, the mass that escapes the binary is modeled as being lost from the vicinity of the accretor through a fast wind, as described by Tauris & van den Heuvel (2006, their “$\beta$” parameter).

Our MESA models are summarized in Table 2. Figure 7 shows the evolution of 4 models in $T_{\text{eff}} - \log g$ space. The left panel shows a model with initial masses $M_1 = 4 M_\odot$, $M_2 = 3 M_\odot$, and an initial period of 2.3 days. The components of the binary evolve normally until the end of the primary’s main sequence lifetime. Shortly before the primary crosses the Hertzsprung gap, it overflows its Roche lobe and begins stable mass transfer to the companion. The primary’s envelope expands further as more mass is removed, and the star moves up the red giant branch. At the tip of the red giant branch (TRGB), the temperatures required to maintain hydrogen burning can only be maintained if the envelope contracts. At this point, when the primary’s luminosity is of order $1000 L_\odot$, the primary leaves the RGB and passes through the region of $T_{\text{eff}} - \log g$ space where HR 6819 and LB-1 are currently observed. The star then continues to contract and heat until it reaches $T_{\text{eff}} \approx 30 kK$ and $\log g \approx 5.5$, where it settles as a core helium burning hot subdwarf (sdOB star). The core helium burning phase would last of order $10^8$ years, but for most of the models we consider, it is truncated prematurely when the secondary (the Be star) leaves the main sequence and overflows its Roche lobe.

The right panel compares stripped star models with masses between 0.49 and 0.83 $M_\odot$. These are produced from calculations...
with initial primary masses of 3.5, 4.5, 5.5, and 7 $M_\odot$. We reduce the efficiency of mass transfer as the primary mass increases, so that the mass of the secondary during the stripped star phase is always about 6 $M_\odot$. As the total mass of the core increases, its radius and luminosity increase, and $\log g$ decreases. However, the sensitivity of these parameters to mass is relatively weak in this region of parameter space, so HR 6819’s atmospheric parameters are consistent with models with a broad range of core masses (and hence initial masses).

Figure 8 shows the evolution of the model with $M_1 = 4M_\odot$ and $M_2 = 3M_\odot$ after the primary begins to contract. We designate the “bloated stripped star” phase of the binary’s evolution as the phase after the tip of the red giant branch when the primary (the B star) has $\log g$ between 2 and 3.5. This is roughly the period when the star passes through the inset in Figure 7, and when its luminosity is first dominated by He burning. The “bloated stripped phase lifetime” is the time it takes the primary to contract from $\log g = 2$ to $\log g = 3.5$ as it moves from the TRGB toward the extreme horizontal branch.

Table 2. Summary of our MESA binary models.

| $M_{1,\text{init}}$ | $M_{2,\text{init}}$ | $P_{\text{init}}$ | $f_{\text{init}}$ | $P_{\text{stripped}}$ | $M_{\text{stripped}}$ | $M_{\text{Be,core}}$ | $M_{\text{env}}$ | $M_{\text{Be}}$ | $t_{\text{RLOF}}$ | $t_{\text{TRGB}}$ | $t_{\text{dOB}}$ | bloated stripped phase lifetime [10$^7$ yr] |
|---------------------|---------------------|-------------------|-------------------|-----------------------|-----------------------|---------------------|----------------|----------------|----------------|----------------|----------------|--------------------------------|
| 3.5                 | 2.8                 | 2.3               | 1                 | 90                    | 0.49                  | 0.43                | 0.06           | 5.8            | 232            | 269.9          | 272.7          | 2.07                                          |
| 4.0                 | 3.0                 | 2.3               | 1                 | 91                    | 0.54                  | 0.47                | 0.07           | 6.4            | 162            | 203.8          | 205.5          | 1.29                                           |
| 5.5                 | 4.2                 | 2.3               | 0.3               | 152                   | 0.67                  | 0.11                | 0.56           | 5.6            | 71             | 97.1           | 98.7           | 2.20                                          |
| 7.0                 | 4.0                 | 2.3               | 0.2               | 79                    | 0.83                  | 0.15                | 0.68           | 5.2            | 41             | 55.8           | 57.2           | 2.27                                          |

$M_{1,\text{init}}, M_{2,\text{init}},$ and $P_{\text{init}}$ are the initial masses and period at the zero-age main sequence. $f_{\text{init}}$ is the efficiency of mass transfer, defined such that a fraction $f_{\text{init}}$ of mass lost by the donor is gained by the accretor. $P_{\text{stripped}}$ is the orbital period during the bloated stripped star phase, when $2 < \log g < 3.5$ (roughly when the star passes through the inset in Figure 7). $M_{\text{stripped}}, M_{\text{Be,core}},$ and $M_{\text{env}}$, are the total, core, and envelope mass of the initial primary (the B star) during this period, and $M_{\text{Be}}$ is the mass of the companion during this period. $t_{\text{RLOF}}, t_{\text{TRGB}},$ and $t_{\text{dOB}}$ are the system age when the primary first overflows its Roche lobe, when it reaches the tip of the red giant branch, and when its luminosity is first dominate by He burning. The “bloated stripped phase lifetime” is the time it takes the primary to contract from $\log g = 2$ to $\log g = 3.5$ as it moves from the TRGB toward the extreme horizontal branch.
the primary has the temperature of a B star and is more inflated than a main sequence star of the same temperature. For the 4 models shown in Figure 7, the duration of this phase ranges from 1.3 to 2.3 ×10^7 years. The secondary (i.e., the Be star) does not evolve on the timescale shown in Figure 8: for all our models, the mass transfer timescale near the TRGB is significantly longer than the thermal timescale of the secondary, so the secondary remains in thermal equilibrium.

We note that, contrary to what is sometimes assumed in the literature (e.g. Eldridge et al. 2020), the contraction from the red giant branch to the extreme horizontal branch does not occur on a Kelvin-Helmholtz timescale of the donor star–the bottom panel of Figure 8 shows that it is longer by more than a factor of 10. The reason for this is that gravitational contraction is not the dominant source of energy during this phase; hydrogen shell burning and core helium burning both contribute luminosity to slow the star’s contraction.

The final period produced by these calculations is somewhat longer (P = 79 – 152 days) than the observed P = 40.3 days in HR 6819. This can be understood as a result of the orbit’s expansion during mass transfer. When mass transfer is conservative, the final period can be calculated analytically to be

\[
P_f = P_i \left( \frac{M_{\text{donor},i}}{M_{\text{donor},f}} \times \frac{M_{\text{accretor},i}}{M_{\text{accretor},f}} \right)^3.
\]

Here \(P_i\) and \(P_f\) represent the initial and final period, \(M_{\text{donor},i}\) and \(M_{\text{donor},f}\) the initial and final masses of the donor, and \(M_{\text{accretor},i}\) and \(M_{\text{accretor},f}\) the initial and final masses of the accretor. In our calculations \(M_{\text{donor},i}/M_{\text{donor},f} = 7 – 8.5\), while \(M_{\text{accretor},i}/M_{\text{accretor},f} = 1 – 2.1\). This causes the final period to be a factor of \(\geq 40\) larger than the initial period. The final period according to Equation 4 could be shorter if the initial mass ratio were more unequal (\(M_1 \gg M_2\)), but in this case mass transfer eventually becomes dynamically unstable, leading to an episode of common envelope evolution (CEE) and typically inspiral to a much shorter period (e.g. Heber 2016).

No analytic expression for the final period exists for non-conservative mass transfer. How the angular momentum of a binary changes following mass loss depends on how the mass is lost (see Soberman et al. 1997, and Tauris & van den Heuvel 2006, their Equation 16.18). In MESAbinary, it is possible to specify what fraction of transferred mass is lost from the vicinity of the donor, from the vicinity of the accretor, and from a circumbinary toroid. Particularly near the end of the stripping process, when \(M_{\text{donor}}/M_{\text{accretor}} \ll 1\), the orbital response to mass loss, which is parameterized through the Tauris & van den Heuvel (2006) subgrid model, depends quite sensitively on the assumed mass loss fractions for each channel. For our calculations with non-conservative mass transfer, the final period decreases by up to a factor of 10 if we assume mass is lost from a circumbinary toroid rather than from a wind in the vicinity of the accretor. Because we have limited physical intuition for what the efficiency of mass transfer is in this problem or how mass that leaves the binary escapes, the observed period cannot reliably constrain the binary’s initial period.

We note that while luminosity and radius increase monotonically with the total mass of the stripped star in our calculations (right panel of Figure 7), this is not necessarily expected to generically be the case. The radius and luminosity of the stripped star depend on the relative masses of the envelope and the He core, with lower envelope masses corresponding to larger radii and luminosities at fixed temperature. For example, Driebe et al. (1998) calculated evolutionary tracks for low-mass stripped pre-He WDs formed from a 1M⊙ progenitor by artificially enhanced mass loss on the RGB. Their 0.41M⊙ track is nearly coincident with our 0.83M⊙ track, while their 0.33M⊙ track falls below our 0.49M⊙ track and is consistent with the LB-1 atmospheric parameters (see also Irgang et al. 2020). In HR 6819, both the dynamical mass and atmospheric parameters of the stripped star are consistent with a scenario in which...
it is a pre-He white dwarf, with a degenerate core and too low a mass to ever ignite helium. However, our binary evolution calculations do not produce stripped stars with such low mass. Because the stripped star must have initially been the more massive component of the binary, and the total initial mass must have been at least $3M_\odot$, the initial mass of the stripped star must have been at least $6M_\odot$. Stars with initial masses near $3M_\odot$ can produce stripped He stars with masses of order $0.4M_\odot$, but these ignite He burning on the RGB. Because their cores are not degenerate during the bloated stripped phase, they have thicker hydrogen envelopes, higher log $g$, and lower luminosity than the Driebe et al. (1998) models.

We have not performed an exhaustive search of parameter space with detailed binary evolution calculations, so it is possible that another region of parameter space could produce systems similar to HR 6819 and LB-1, even though no other satisfactory models were found in the BPASS library. Below, we briefly summarize how the MESA models change when the initial conditions or model parameters are perturbed.

- **Longer periods**: For longer initial periods, mass transfer begins later, when the primary has expanded more. For $P_{\text{init}} \lesssim 20$ days, evolution is similar to our fiducial models, but stripping occurs over a shorter timescale, the final period is longer, and a somewhat larger fraction on the envelope remains when contraction begins, shifting models downward by up to $\approx 0.2$ dex in Figure 7.

  For $20 \lesssim P_{\text{init}}/\text{days} < 150$, mass transfer begins when the envelope is convective and the core mass is too small to support stable mass transfer, leading to dynamical instability and presumably an episode of CEE. MESA does not include robust treatment of CEE, but it is expected to result in shorter periods than the observed $P = 40.3$ days (e.g. Ivanova et al. 2013).

  For $150 \lesssim P_{\text{init}}/\text{days} < 2000$, the primary begins core He burning before RLOF. Stripping then begins on the asymptotic giant branch (AGB), where the larger core mass allows mass transfer to remain dynamically stable (e.g. Chen & Han 2008). Eventually, the primary is reduced to a carbon-oxygen core with a thin helium envelope, whose contraction is similar to that of a single proto-white dwarf. These models pass through the same region of $T_{\text{eff}} - \log g$ space highlighted in Figure 7, typically shifted upward somewhat from our fiducial models of the same total mass. However, they contract much more rapidly than our fiducial models, with lifetimes of $\lesssim 10^3$ years. The final periods are also much longer, $P \gtrsim 3000$ days, though we note that these depend sensitively on the assumptions made about how mass is lost from the binary.

- **Shorter periods**: For shorter periods, RLOF occurs earlier, when the primary is still on the main sequence. For $P_{\text{init}} \lesssim 2$ days, the primary loses about half its mass during the initial phase of RLOF, but then essentially continues its evolution as a lower-mass main sequence star. Meanwhile, the now-more-massive secondary evolves faster than the primary, reaching the end of the main sequence before the primary.

- **More unequal mass ratios**: If the initial masses of the two components are too unequal ($M_2/M_1 \lesssim q_{\text{crit}}$, where $q_{\text{crit}}$ depends on the structure of the donor; e.g. Hjellming & Webbink 1987; Han et al. 2002; Ge et al. 2010), mass transfer eventually becomes dynamically unstable, leading to an episode of CEE. When mass transfer begins, the envelope is radiative, and mass transfer is stable for a wide range of mass ratios. However, the envelope becomes convective soon after mass transfer begins, and mass transfer then becomes unstable. We note that the secondary is generally not expected to accrete much mass during CEE; this also speaks against an initially low secondary mass.

- **Higher primary masses**: As the mass of the primary increases, the efficiency of mass transfer must approach 0 to avoid making the Be star too massive. If this is achievable physically, then a higher primary mass produces leads to a more massive stripped He star. For our default parameters, $M_{1,\text{init}} = 10M_\odot$ leads to a 1.3 $M_\odot$ stripped He star. During the contraction phase, the 1.3$M_\odot$ model continues...
Increasing the larger the overshooting efficiency parameter, the larger the changes the mapping between initial mass and stripped star mass. Figure 7, increasing modelstoward lower so the observed mass of the Be star cannot be reproduced. thetotalinitialmassofthebinaryislessthanitscurrenttotalmass, consistent solution between the dynamically-inferred mass of the report a dynamical mass of \( M \approx \). We note that while the models shown in Figure 7 allow for a \( M > 1 \) order \( 1 \) higher than this. A 1.5\( M_\odot \) He star is generically difficult to produce in the context of our models (or the BPASS models); if stripping occurs near the Hertzsprung gap, it requires a primary mass of order 12\( M_\odot \) for our adopted overshooting parameter. Unless mass transfer is highly inefficient, the Be star would end up too massive in such a scenario. Without binary interactions, a single star of solar metallicity must have initial mass \( \gtrsim 9 M_\odot \) to reach a He core mass of 1.5\( M_\odot \) prior to core He burning.

The dynamical mass constraint derived by Shenar et al. (2020) comes primarily from the Hz line. Since spectral disentangling relies on the assumption that the individual spectra of both components are time-invariant, and this assumption likely does not hold in detail for the Hz line of LB-1 due to phase-dependent disk irradiation by the B star (Liu et al. 2019), it is possible that the dynamical mass measurement of the B star is biased and the true mass is lower. If the mass measurement is reliable, it would imply a significantly higher envelope mass (\( M_{\text{env}}/M_{\text{He core}} \sim 1 \)) than produced by our models, since the luminosity during the stripped phase is set mainly by the He core mass. Three of the five confirmed Be + sdOB binaries have estimated masses above 1.0\( M_\odot \) (Wang et al. 2017), so a higher stripped-star mass is not beyond the realm of possibility. This is, at least partially a selection effect, since higher-mass sdOBs are hotter and easier to detect (e.g. de Mink et al. 2014; Wang et al. 2017; Schootemeijer et al. 2018). We also note that the secondaries in these systems have higher masses than in HR 6819 (9 – 12\( M_\odot \)), allowing for higher initial masses for the primaries.

3 SUMMARY AND DISCUSSION

We have shown that HR 6819 is a binary containing a rapidly rotating Be star and an undermassive B star with mass of order 0.5\( M_\odot \). Radial velocity measurements from double-lined spectra establish that the two luminous components of the binary are orbiting one another (Figure 4). There is thus no need to invoke an unseen 3rd object to explain the orbital motion of the B star, as proposed by Rivinius et al. (2020). We propose that the B star is the core of a star with initial mass of order 5\( M_\odot \) whose envelope was recently stripped by its companion near the tip of the red giant branch. In this scenario, the B star is currently in a short-lived (few \( \times 10^7 \) years) phase of contraction toward the extreme horizontal branch, where it will become a core helium burning sdOB star (e.g. Heber 2016). In this scenario, HR 6819 is a progenitor system for Be + sdOB binaries, of which at least five are known (e.g. Gies et al. 1998; Koubysky et al. 2012; Peters et al. 2013, 2016; Wang et al. 2017; Dulaney et al. 2017) and another dozen candidates have been identified (Wang et al. 2018). Recent accretion of the B star’s envelope would then explain the near-critical rotation of the Be star.

HR 6819 is quite similar to the binary LB-1 (Liu et al. 2019), which has recently been the subject of much discussion. In El-Badry & Quataert (2020), we speculated that the emission lines in LB-1 might come from a circumbinary disk around a BH and a stellar companion. But spectral disentangling by Shenar et al. (2020) has since revealed the presence of a Be star in the spectrum, similar to the case in HR 6819. The B star in LB-1 would then also be a recently stripped He star, with no BH in the system.

Further modeling will be required to pin down the mass and current evolutionary state of these systems. Our MESA calculations produce stripped stars with masses, temperatures, and surface gravities comparable to HR 6819 and LB-1, but there is not yet full agreement between the dynamically- and evolutionarily-inferred masses for LB-1. For both systems, Gaia DR3 astrometry and ground-based interferometry are expected to prove useful in further constraining the system parameters.

3.1 Implications for the formation of Be stars

We now consider what the existence of the HR 6819 system implies for the broader Be star population.

HR 6819 is bright enough (\( G = 5.22, V = 5.36 \)) to be seen with the naked eye. Estimating the rate of Be + stripped star binaries based on HR 6819 is challenging both because the search that led to its discovery does not have a well-defined selection function and because Poisson errors are large when only one object is known. We nonetheless make a crude estimate of the frequency of HR 6819-like binaries below.

Roughly 15% of all field B stars are Be stars (Jaschek & Jaschek 1983; Zorec & Briot 1997). There are about 1500 B stars brighter than \( G = 6 \), corresponding to about 225 Be stars. Most stars this bright have been studied in some detail, so it seems a reasonable assumption that HR 6819 is the brightest object in its class, and that there are at most a few stars brighter than \( G = 6 \) in its current evolutionary state.

Once the stripped star contracts, HR 6819 will likely continue to appear as a Be star for a few \( \times 10^7 \) years – a few hundred times the lifetime of the stripped star phase – depending on the mass of the Be star. This implies that for every system like HR 6819, there should be a few hundred binaries consisting of a Be star and a faint sdOB or WD companion. Such binaries certainly exist, but they are challenging to detect because the sdOB typically contributes at tiny fraction of the light in the optical. All 5 of the confirmed Be + sdOB binaries are brighter than \( G = 6 \) and were discovered via UV spectra. Recently, Wang et al. (2018) identified another dozen

\(^2\) We estimate the number of B stars brighter than \( G = 6 \) based on Gaia photometry. There are 1275 sources in Gaia DR2 with \(-0.3 < \text{G}_{\text{BP}}-\text{G}_{\text{RP}} < 0.1\), a color range corresponding to \( T_{\text{eff}} = (11 - 30) \) kK without extinction. We inflate this to 1500 to account for extinction (which is modest for most naked-eye stars) and incompleteness in Gaia DR2 for bright stars. We obtain a similar estimate based on the Hipparcos catalog. The Be star catalog from Jaschek & Egret (1982) contains 141 sources brighter than \( V = 6 \) but does not claim to be complete.
candidate Be + sdOB systems using IUE spectra, all of which are brighter than $G = 7$. Schootemeijer et al. (2018) estimated that there should be 30-50 yet undiscovered Be + sdOB binaries for each confirmed system of a given magnitude, which corresponds to 150-250 Be+sdOB binaries brighter than $G = 6$. Raguëza (2001) estimated through population synthesis simulations that about 70% of Be stars formed through binary evolution should have faint white dwarf companions, while 20% should have sdOB companions. No white dwarf companions have yet been unambiguously detected.

Given our estimate that there are only a few hundred Be stars brighter than $G = 6$, the existence of HR 6819 is consistent with a scenario in which a majority of Be stars form through this evolutionary channel. This is an appealing possibility, because the origin of the Be phenomenon is still not fully understood. It is certainly tied to rapid rotation, but there is considerable debate about how Be stars were spun up in the first place (e.g. Rivinius et al. 2013). Accretion of mass from a companion is one possibility. There is, however, considerable uncertainty in the frequency of HR 6819-like binaries. Given $N = 1$ objects discovered among 225 Be stars, the 1-sigma Poisson confidence interval for the fraction of Be stars in this evolutionary phase is (0.08 – 1.5%). If we take the mean lifetime of a Be star to be $5 \times 10^7$ years and the mean lifetime of the stripped star to be $2 \times 10^5$ years, the lower limit of 0.08% implies that at least 20% of Be stars went through a phase like HR 6819.

HR 6819 and LB-1 are currently the only known objects in their class. Given that LB-1 is more than 6 magnitudes fainter than HR 6819 and was discovered in an multi-epoch RV survey of only 3000 objects (without selection on color or spectral type), there are probably significantly more similar systems with magnitudes between HR 6819 and LB-1 that have yet to be discovered. We estimate that there are of order $10^4$ Be stars brighter than $G = 12$. Given the rate estimate from HR 6819, this implies between 8 and 150 Be + bloated stripped star binaries brighter than $G = 12$, a large fraction of which should be discoverable to wide-field spectroscopic surveys such as SDSS-V. While both HR 6819 and LB-1 were only identified as candidate stripped stars after significant spectroscopic study, additional candidates could likely be identified from a single medium-resolution spectrum as objects exhibiting both the emission lines of a classical Be star and narrow absorption lines from the companion (indicative of low $v \sin i$). Normal Be stars viewed pole-on would be a contaminant to such a search, but these can usually be identified from the shape of their emission lines (Hummel & Vrancken 2000; Rivinius et al. 2013).

With LB-1 and HR 6819 demoted, the list of detached stellar mass BH candidates remains short. The most secure detached BH candidates to date are likely those discovered in globular clusters with MUSE (Gierschner et al. 2018, 2019), but the evolutionary history of these objects is likely quite different from that of BH binaries in the field. No luminous companion has been discovered to the giant-BH binary candidate reported by Thompson et al. (2019), but its status as a BH is not uncontroversial (van den Heuvel & Tauris 2020; Thompson et al. 2020). We note that the other two BH candidates on the field also contain Be stars (Casares et al. 2014; Khokhlov et al. 2018). These systems are not likely to contain low-mass stripped stars, because in them it is the Be star that is clearly RV variable. Be stars do not constitute a dominant fraction of all massive stars, and there is no strong bias toward Be stars in the selection function of multi-epoch RV surveys. If these objects do contain BHs, then their status as Be stars is likely to be a result of accretion from the BH progenitor.

In the final stages of this manuscript’s preparation, a similar analysis of HR 6819 was completed by Bodensteiner et al. (2020). Overall, there is excellent agreement between our results.

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APPENDIX A: STELLAR PARAMETERS ANCHORED TO EVOLUTIONARY TRACKS AND DISTANCE

Figure A1 compares the atmospheric parameters of both components of HR 6819 compared to MIST evolutionary tracks for normal (i.e. non-stripped) stars. If the B star were a normal star, its $T_{\text{eff}}$ and $\log g$ would imply a mass of at least $9 M_{\odot}$, roughly 20 times the mass we infer dynamically. For the Be star, the atmospheric parameters imply a mass between 5 and $8 M_{\odot}$. 

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make no attempt to account for the different IMF probability or lifetime of each track (which would downweight higher-mass tracks), or the observational selection function (which would downweight lower-mass tracks). We then retain samples with probability proportional to the likelihood of their $T_{\text{eff}}$ and log $g$ given the measured constraints on the Be star’s atmospheric parameters, approximating the likelihood function as a two-sided Gaussian. This effectively selects points from the evolutionary tracks that are close to the Be star in $T_{\text{eff}}$ – log $g$ space. We use the resulting samples to estimate the mass, radius, and luminosity of the Be star, and to propagate forward the uncertainty in other parameters in Table 1 that depend on the Be star’s mass.

The system parameters listed in Table 1 are anchored to the mass of the Be star estimated from its $T_{\text{eff}}$ and log $g$. The constraints can be tightened somewhat if the distance, $d = (310 \pm 60)$ pc, is taken into account. To this end, we add a term to the likelihood function that compares the predicted V-band apparent magnitude of the evolutionary tracks to the measured $V = 5.36$, assuming a V-band flux ratio of $f_{\text{Bu}}/f_{\text{tot}} = 0.52 \pm 0.07$ and extinction $A_V = 0.42$. The resulting constraints are listed in Table A1. They are consistent with the constraints in Table 1, but the uncertainties are somewhat smaller. At fixed $T_{\text{eff}}$, a constraint on luminosity translates directly to a constraint on radius. When combined with the measured log $g$, this translates to an improved constraint on mass.

### A1 System parameters independent of evolutionary tracks

An estimate of the stellar parameters of both components that is independent of evolutionary tracks can be calculated from the apparent magnitude, distance, and flux ratio. The V-band apparent magnitude of HR 6819 is $5.36$. Given extinction $A_V = 0.42$ mag, total bolometric correction BC $= -1.5 \pm 0.15$ mag and distance $d = 310 \pm 60$ pc, the absolute bolometric magnitude of HR 6819 is $M_{\text{bol}} = -4.05^{+0.4}_ {-0.5}$. This corresponds to a bolometric luminosity $L_{\text{bol,tot}} = 10^{0.4(M_{\text{bol,tot}}-M_{\odot})} L_{\odot}$ (A1) $= 3200^{+1500}_ {-1200} L_{\odot}$, (A2)

where $M_{\text{bol,tot}} = 4.74$. Given the effective temperatures of the two components, their respective uncertainties, and $(47 \pm 5)$% flux contribution from the Be star at $J = 4000 \text{ Å}$, the fraction of the total light contributed by the Be star is $f_{\text{Bu}}/f_{\text{tot}} = 0.43^{+0.08}_ {-0.07}$. The individual luminosities are then:

$$ \log(L_{\text{bol,B}}/L_{\odot}) = 3.13^{+0.18}_ {-0.21} \quad \text{(A3)} $$

$$ \log(L_{\text{bol,Be}}/L_{\odot}) = 3.25^{+0.18}_ {-0.20} \quad \text{(A4)} $$

The radius of each component can then be calculated as $R = \sqrt{L_{\text{bol}}/(4\pi \sigma_{\text{SB}} T_{\text{eff}}^4)}$, where $\sigma_{\text{SB}}$ is the Stefan Boltzmann constant. This yields

$$ R_{\text{B}} = 4.8^{+1.1}_ {-1.2} \ R_{\odot} \quad \text{(A5)} $$

$$ R_{\text{Be}} = 4.5^{+1.8}_ {-1.2} \ R_{\odot} \quad \text{(A6)} $$

Finally, the mass of each component can be calculated as $M = g R^2 / G$, yielding

$$ M_{\text{B}} = 0.46^{+0.71}_ {-0.28} M_{\odot} \quad \text{(A7)} $$

$$ M_{\text{Be}} = 4.8^{+1.6}_ {-3.0} M_{\odot} \quad \text{(A8)} $$

Equations A3-A8 are fully consistent with the constraints in Tables 1 and A1, which are anchored on the fit of the temperature and gravity of the Be star with evolutionary tracks. Equations A3-A8 are also consistent with, but independent of, the dynamically measured mass ratio. We adopt the parameters anchored to evolutionary tracks as our fiducial values due to their tighter constraint on the component masses.

### APPENDIX B: SPECTRAL DISENTANGLING AND DETERMINATION OF $K_{\text{BE}}$

Spectral disentangling solves for the two single-star spectra that can best reproduce all the observed composite spectra for an assumed set of relative RVs at each epoch. If the RVs are not known a priori, they can be solved for in tandem with the single-star spectra by searching for the RVs that allow the disentangled single-star spectra to best reproduce all the observed composite spectra. In the case of HR 6819, the problem is simplified by the fact that the narrow lines of the B star allow a preliminary estimate of its RVs to be made directly from the composite spectra.

We optimize for the RVs in two steps. In the first step, we assume that the velocity of the B star at each epoch is the velocity predicted by the orbital solution of Rivinius et al. (2020). We then step through a grid of RV semi-amplitudes for the Be star, $K_{\text{B,Be}}$, calculating the optimal disentangled spectra for each value. The choice of $K_{\text{B,Be}} = 0 \text{ km s}^{-1}$ corresponds to a stationary Be star, as would be expected in a triple scenario.\(^5\) We choose the optimal $K_{\text{B,Be}}$ and corresponding spectral decomposition as the value that minimizes the total $\chi^2$ summed over all observed spectra.

With the disentangled spectra from the first step in hand, we use them as templates to fit the composite spectra for the RVs at each epoch, as described in Section 2.4. This revealed intrinsic scatter of $s \approx 4 \text{ km s}^{-1}$ for the RVs of the B star (see Section 2.4).

In the second step, we repeat the disentangling, again stepping through a grid of $K_{\text{B,Be}}$ values. Rather than setting the RVs of the B star to the values predicted by the Rivinius et al. (2020) orbital solution, we now set them to the values measured in the previous step. The best-fit values of $K_{\text{B,Be}}$ for a given spectral region do not

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\(^4\) For solar-metallicity MIST models, $T_{\text{eff}}$ values of 16 kK and 18 kK correspond to bolometric corrections of $-1.35$ and $-1.64$, with weak dependence on log $g$.

\(^5\) Note that there is no need to try different center-of-mass velocities: because we do not use any template spectra, the absolute RV scale is arbitrary.
change much between the first and second step, but the quality of the fit improves somewhat in the second step.

We perform spectral disentangling and determine the optimal $K_{\text{Be}}$ for 10 different regions of the spectrum, each centered on an emission or absorption line with clear contributions from both components. Treating different lines independently allows us to test whether different ions trace different kinematic components, and to quantify the uncertainty in the $K_{\text{Be}}$ values we infer.

Figure B1 shows the total $\chi^2$ as a function of $K_{\text{Be}}$ for the 10 lines. For easy comparison, we subtract the minimum $\chi^2$ for each line. The Fe II lines in the top panel are dominated by emission in the Be star spectrum (e.g. Figure 6). The He I lines (middle panel) are dominated by absorption. The Balmer lines in the Be star spectrum (e.g. Figure 6). The He I lines (middle panel) are dominated by absorption. The Fe II lines in the top panel are dominated by emission.

The 10 spectral regions we consider have optimal $K_{\text{Be}}$ ranging from 2 to 10 km s$^{-1}$. Of these, 7 have optimal values between 3 and 6 km s$^{-1}$. None of the lines have minimal $\chi^2$ values associated with $K_{\text{Be}} = 0$ km s$^{-1}$, as would be expected for a stationary Be star. The scatter in $K_{\text{Be}}$ values inferred from different lines, including lines tracing the same ion, is larger than would be expected if the formal uncertainties in $K_{\text{Be}}$, as quantified by $\chi^2(K_{\text{Be}})$ for a single line, were reliable. This is most likely a consequence of small errors in continuum normalization as well as time-variability in the flux ratio, which violate the underlying assumptions of the spectral disentangling algorithm (see Hensberge et al. 2008).

We do not find systematic difference between lines from different ions, or emission vs. absorption lines. This suggests that the emission lines, which are presumed to originate in a disk around the Be star, move coherently with the star. Considering the 10 spectral regions shown in Figure B1, as well as the RVs obtained when fitting a synthetic spectral template (Figure B3), we adopt $K_{\text{Be}} = (4.5 \pm 2)$ km s$^{-1}$ as our fiducial value. We then fix $K_{\text{Be}} = 4.5$ km s$^{-1}$ in disentangling the full spectrum.

The shape of the disentangled spectra depends very weakly on the value of $K_{\text{Be}}$. Choosing $K_{\text{Be}} = 0$ km s$^{-1}$ makes both the emission and absorption lines of the disentangled Be star spectrum slightly broader (by an order of 0.1 A) than in our fiducial solution. Broader lines suppress the RV variability of the Be star, but this leads to a worse overall fit to the composite spectra. The inferred $T_{\text{eff}}$ and $\log g$ of the two components are not sensitive to the adopted value of $K_{\text{Be}}$, at least over the range $0 \leq K_{\text{Be}}$/km s$^{-1} < 10$.

In addition to our primary spectral decomposition using CRES, we also experimented with using the Fourier-space spectral disentangling code FDBinary (also referred to as FDBinary in the literature; Ilijić et al. 2004; Ilijić 2017), which uses the algorithm described by Hadrava (1995) to separate composite spectra. The disentangled spectra produced by the two codes are very similar. Figure B2 compares the two codes’ constraints on $K_{\text{Be}}$ from the H$\beta$ line. Under idealized circumstances, the velocity-space and Fourier-space spectra disentangling algorithms are mathematically identical (e.g. Hensberge et al. 2008). The most significant practical difference in the implementations for our purposes is that FDBinary requires the RVs of both components to follow Keplerian orbits, while CRES allows them to be specified manually at each epoch. The FDBinary solution thus does not model the scatter in the Be star’s RVs, which is accounted for in the CRES solution.

We also experimented with fitting RVs for both components.
using synthetic spectral templates. Figure B3 shows the best-fit RVs of the Be star obtained by jointly fitting the RVs of both components at each epoch, using the best-fit TLUSTY/SYNSPEC spectra for both components and fitting the He I absorption line at 4387.9 Å. Because the synthetic spectra do not perfectly reproduce the observed spectra, the narrow lines in the Be star’s spectrum are washed out by rotation, and many of the Be star’s lines are significantly contaminated by emission, we find that we obtain more stable RVs when fitting a narrow wavelength range than when fitting the full spectrum simultaneously. The RVs measured this way show larger scatter than those measured from the disentangled spectra (Figure 4), but they still show clear periodic modulation, the amplitude of which is consistent with the amplitudes we measure from spectral disentangling. The RVs we measure for the B star are in good agreement with those measured in Section 2.4.

APPENDIX C: B STAR VARIABILITY

As discussed in Section 2.4, the measured RVs of the B star do not perfectly follow the predictions of a Keplerian orbit, but exhibit additional scatter with amplitude $\sigma_{RV} \approx 4 \text{ km s}^{-1}$. Here we show that this variability is accompanied by variability in the B star’s line profiles.

The top panel of Figure C1 shows the line profiles of two Si III lines from three single epoch spectra of HR 6819, which are chosen to showcase the range of profiles found in the dataset. Weak and narrow metal lines such as these are ideal for characterizing the B star’s variability because contamination from the Be star – besides dilution by its continuum – is negligible. All three epochs are shifted to rest frame. The depth and width of the lines vary visibly between epochs. We quantify the line width with $v_{\text{macro}} = \sqrt{2} \sigma$, where $\sigma$ is the dispersion of the best-fit Gaussian broadening profile (see Appendix D).

The three lower panels of Figure C1 show $v_{\text{macro}}$ equivalent width, and RV deviation from a Keplerian orbit, measured from the Si III 4568 line across all 51 epochs. The epoch-to-epoch variation in $v_{\text{macro}}$ is of order 10%, with a maximum-to-minimum range of almost a factor of 2. The typical variation in equivalent width is somewhat smaller, of order 5%. The RV scatter is 4 km s$^{-1}$, consistent with our finding in Section 2.4. There is no clear trend with orbital phase. Like the photometric variability (see Rivinius et al. 2020), the line profile variability exhibits power on timescales from $\approx 0.4$ days to several days, and is not dominated by a single period.

The observed line profile variations cannot be understood as a result of photometric variability of the Be star, which would change the flux ratio – and thus, line depth and equivalent width – but not $v_{\text{macro}}$. It thus seems natural to connect the B star’s line profile variations with the observed photometric variability of HR 6819. The period and amplitude of the photometric variability are con-
sistent with a scenario in which they are due primarily to the B star (see below). We note, however, that photometric variability is also common among Be stars, so we cannot rule out the possibility that some of the observed photometric variability is due to the Be component.

The period, photometric variability amplitude, RV scatter, temperature, and luminosity of the B star in HR 6819 are all typical of slowly pulsating B-type (SPB) stars (e.g. Waelkens et al. 1998). Indeed, the prototypical SPB star, 53 Persei, has $T_{\text{eff}} = 17$ kK, $\log(L/L_\odot) = 2.9$, and exhibits pulsation-driven RV variations with peak-to-peak amplitude of about 15 km s$^{-1}$, and photometric peak-to-peak amplitude of order 0.2 mag, quite similar to the B star in HR 6819 (Smith et al. 1984; Chapellier et al. 1998). SPB stars undergo non-radial g-mode pulsations, which are excited by the $\kappa$ mechanism acting on the metal opacity bump for $T_{\text{eff}} = 10 - 25$ kK (e.g. Dziembowski et al. 1993). Line profile variations in SPB stars are primarily due to variations in the velocity field at the stellar photosphere; variations due to changes in $T_{\text{eff}}$ are usually negligible (Buta & Smith 1979; Townsend 1997; Aerts & Eyer 2000).

Photometric and spectral variability violate the assumptions of the spectral disentangling algorithm. When variability is purely photometric, it is straightforward to account for a time-dependent luminosity ratio (e.g. Ilijic 2004; Hensberge et al. 2008). In the presence of line profile variations, however, there is no unique spectral decomposition, so the disentangled spectra represent a time-average of the component spectra. We carried out tests with mock data to assess the effects of unmodeled line profile variability similar to that found in HR 6819 on spectral disentangling. Because the variability is not systematically phase-dependent, we find that it does not lead to systematic biases in $K_{\text{B}}$, or in the width or depths of the disentangled line profiles, but it does reduce the total goodness-of-fit.

### APPENDIX D: ROTATION VS. MACROTURBULENT BROADENING

The metal lines of the B star component of HR 6819 have a typical FWHM of 50 km s$^{-1}$. This is relatively narrow for a B star, but it is larger than the expected thermal width given the B star’s spectral type in the absence of rotation or macroturbulent broadening (FWHM $\approx 10$ km s$^{-1}$).

Figure D1 shows line profiles of the Si III 4568 and 4575 Å lines in a single-epoch ($\phi \approx 0.55$) spectrum of HR 6819. These lines are ideal for characterizing the B star’s line broadening because they are intrinsically narrow and spectral disentangling suggests that the contribution of the Be star to them is negligible. We compare the observed line profiles to the predictions of a TLUSTY/SYNSPEC model ($T_{\text{eff}} = 16$ kK and $\log g = 2.75$) with no rotational or macroturbulent broadening, and to line profiles broadened only by macroturbulence or rotation. The rotational broadening profile is taken from Gray (1992), using a linear limb darkening coefficient $\epsilon = 0.5$. The macroturbulent broadening profile is modeled as a Gaussian with $\sigma = v_{\text{macro}}/\sqrt{2}$. Rotational broadening alone reproduces the observed line profiles poorly. Disagreement between the observed and modeled profiles is most obvious in the line wings, which for the rotational profiles extend only to $\pm v \sin i$. The observed line wings are less steep and extend farther from the line center. This is not a result of contamination from the Be star: given its $v \sin i \approx 180$ km s$^{-1}$, any absorption from it is spread over a significantly wider wavelength range, $\Delta \lambda > 5$ Å. Purely Gaussian broadening does reproduce the observed line profiles reasonably well. An isotropic Gaussian broadening profile is not well-motivated physically (e.g. Gray 1992; Simón-Díaz & Herrero 2014), but it is a serviceable simplification. The physical basis of non-rotational broadening in B stars is not well understood. As it is modeled here, macroturbulence represents any broadening mechanisms besides rotation, not necessarily turbulence. Given the observed photometric variability and line profile variations in HR 6819, it is likely that pulsation-driven fluctuations in the surface velocity field of the B star contribute to the broadening. Such fluctuations are not expected to produce strictly Gaussian broadening profiles (e.g. Aerts & Eyer 2000; Aerts et al. 2014), but in practice we find that Gaussians provide a reasonably good fit to the line profiles of uncontaminated metal lines in HR 6819 at all epochs, though the best-fit values of $v_{\text{macro}}$ change somewhat from epoch to epoch (Figure C1). The conclusion that rotation alone produces line profiles with too-steep wings also holds at all epochs.

Because non-rotational broadening likely dominates over rotational broadening, we cannot reliably estimate $v \sin i$ of the B star from its line profiles. By fitting the observed Si III 4568 line profiles with a combination of rotational and macroturbulent broadening (e.g. Ryans et al. 2002; Simón-Díaz & Herrero 2014), we can limit the projected rotation velocity to $v \sin i < 20$ km s$^{-1}$. If the B star is tidally synchronized, the limits on its radius and the binary’s inclination (Table 1) would imply $v \sin i = 3.1^{+1.3}_{-1.2}$ km s$^{-1}$.

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Figure D1. Black line shows an observed spectrum of HR 6819, highlighting two Si III lines for which the absorption is dominated by the B star. Blue line shows a model spectrum with no rotational or macroturbulent broadening. Cyan and red lines show fits to the line profiles assuming purely rotational and Gaussian broadening, respectively. The observed line profiles have broader wings than predicted for pure rotational broadening and are reasonably well-fit by Gaussian profiles. This indicates that non-rotational broadening contributes significantly to the observed line widths.