

**The Relationships Between Partial Tidal Disruption Events
and Stellar Evolution in Main Sequence Stars**



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Submitted in fulfillment of the senior thesis requirement of the Department of Physics at
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This work is mine unless otherwise cited.

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First Reader

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Date

Date

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The Relationships Between Partial Tidal Disruption Events and Stellar Evolution in Main Sequence Stars

Shannon E. Bratt-Pfotenbauer

Abstract

In this project, I study close encounters between a $1M_{\odot}$ main sequence star, modeled using MESA, with black holes through smoothed particle hydrodynamic simulations. I consider black holes in the stellar mass range, particularly $M_{\text{bh}} = 30M_{\odot}$, as that is the median mass of black holes involved in collisions with stars, specifically within dense star clusters. I observe and analyze how partial tidal disruption events between these celestial objects affect how the usual process of stellar evolution occurs during the star's orbital period around the black hole, and whether secondary passages around the black hole are distinctly different from primary passages.

Within this paper, I describe the process of creating two dynamic collisions where in each collision, the star has an orbital period around the black hole of 1Myr, and their periapsis distances are the only parameters that differ from one another. I evolve the collision products for a time equal to their orbital periods, put them back into their original collisions, and let them make a second passage around the black hole. I then analyze and compare the chemical compositions of each relevant stellar model, and make conclusions about how partial tidal disruption events influence the compositions of the stars involved in such encounters.

1 Introduction and Background

1.1 Motivation

The effects that a partial tidal disruption event (PTDE) between a main sequence (MS) star and an individual black hole (BH) has on the star's evolution motivates this research. The assumption is that during a PTDE, the less dense, hydrogen-dominated outer layers of the MS star will be stripped away and accreted by the BH, but the helium-rich core of the star will be left largely unharmed. Many studies have been conducted regarding collisions between stars and BHs, but what occurs in a star internally after an encounter with a BH appears to be less discussed in the literature. My goal is to explore two individual cases, where PTDE simulations between a sun-like MS star and a stellar mass BH with differing periapsis distances take place. I aim to understand whether the changes in the structure of the stars significantly impacts how their evolution is or is not affected after being disrupted, as well as the impact on future passages around the BH after their initial interactions.

1.2 Prior Work

Plenty of research exists discussing collisions between stars and black holes of every kind. One study of note is Kiroğlu et al. [1], which focuses on PTDEs of a MS star by intermediate mass BHs where the outcomes of the collisions are analyzed, giving particular weight to the significance of the BH mass on said outcomes. The results ranged from total tidal disruption events (TDEs) to PTDEs for stronger encounters, to small tidal interactions for weaker encounters. Each simulation in Kiroğlu et al. uses an identical MS star, and the star used in my simulations is modeled off of the star used in their work. My stellar model is the same mass as

the Sun, part of the main sequence, and has a metallicity content of 0.001 (a metal in astrophysics is any element heavier than helium). The common method of measuring objects in the universe is by comparing their characteristics with the Sun, for example, masses are measured in solar masses ($M_{\odot}$), where $1M_{\odot} = 1.989 * 10^{30}\text{kg}$, and sizes of objects are measured in solar radii ($R_{\odot}$), where $1R_{\odot} = 6.9551 * 10^8\text{m}$. These values and some other astronomical constants can be found at the end of this paper in Appendix A.

Another relevant study is Kremer et al. [2], which examines the interactions between various MS stars and stellar mass BHs. Both of these studies concentrate primarily on what happens to the celestial bodies depending on specific criteria, like whether a star is ejected on an unbound orbit after a collision, or instead gets gravitationally bound by the BH it collided with. The evolution of the star after a close encounter is not commented on, assuming the star exists after said encounter.

2 Theory

2.1 Gravity

Gravity is the fundamental force that we are most familiar with on a daily basis, and yet the one we seem to take the least note of. It keeps our feet on the ground and the Earth orbiting around the Sun, and is the glue that holds the universe itself together. A surface-level discussion of gravity is warranted, given its role in astrophysical interactions both in the real world and in computer simulations.

2.1.1 Laws of Motion

Isaac Newton formulated his theory of gravity, and his three iconic laws of motion that govern every introductory physics course, in 1665 or 1666. Only the second and third laws are of relevance in this paper. Newton's second law states that

$$\vec{F} = m\vec{a}, \quad (2.1)$$

where $\vec{F}$ is the force acting on an object, m is the mass of the object or the center of mass between two objects, and $\vec{a}$ is the acceleration of the object. Since Equation 2.1 uses the center of mass in calculations with more than one object, it is reserved for calculating the interactions between individual particles inside of a star, not collisions between large celestial objects.

Newton's third law states

$$\vec{F}_{12} = -\vec{F}_{21}, \quad (2.2)$$

where $\vec{F}_{12}$ is the force of object 1 acting on object 2, and $\vec{F}_{21}$ is the force of object 2 acting on object 1. Therefore, the law states each force (or any force, for that matter) has an equal and opposite force to counteract it. Equation 2.2 will be used similarly to Equation 2.1 to calculate the forces on the individual particles making up a star: each particle will exert a force on its neighboring particles, and an equal and opposite force will be applied back on the original particle.

2.1.2 Newton's Law of Universal Gravitation

Another one of Newton's fundamental laws, his universal law of gravitation, is important to mention given its relevance in StarSmasher calculations. StarSmasher will be spoken on in more detail in Section 3.2, but it is the smoothed particle hydrodynamics (SPH) code used to run

the collision simulations between my MS star and BH. The universal law of gravitation is given by

$$\vec{F}_{21} = -G \frac{m_1 m_2}{r_{21}^2} \hat{r}_{21}, \quad (2.3)$$

where $\vec{F}_{21}$ is the gravitational force of object 2 acting on object 1, m_1 is the mass of object 1, m_2 is the mass of object 2, r_{21} is the distance between the objects, $\hat{r}_{21}$ is the unit vector pointing from object 2 to object 1, and G is the universal gravitational constant, $G = 6.67 * 10^{-11} \frac{m^3}{kg*s^2}$. Any “G” variables in future equations will be this constant. The negative sign indicates the force is attractive, and the equation would be the same for object 2 exerting a force on object 1, except $\hat{r}$ would point from object 1 to object 2. StarSmasher uses Equation 2.3 to make calculations for large-scale collisions, like between two stars, or between a star and a BH.

2.2 Stellar Structure and Evolution

Understanding what makes up a star, and stellar evolution, is necessary to make proper and well-informed conclusions about my collision simulations. The accuracy of the simulations also relies on the conditions assigned to the star, i.e. if the star is given an incorrect value for mass, radius, metallicity, etc., the collision simulation will have an incorrect and unexpected result.

2.2.1 Hydrostatic Equilibrium

Hydrostatic equilibrium is what makes a star a star, alongside nuclear fusion. In order for a star to be static and not collapse or explode, a pressure gradient $\frac{dP}{dr}$ must be present to

counteract the force of gravity. The condition for hydrostatic equilibrium for spherically symmetrical objects, assuming accelerations are negligible, is given by

$$\frac{dP}{dr} = -G \frac{M_r \rho}{r^2} = -\rho g, \quad (2.4)$$

where P is the pressure, M_r is the star's "interior" mass (the mass inside a sphere of radius r),

$g = \frac{GM_r}{r^2}$ is the local acceleration of gravity at radius r , and ρ is the density of the star's gas.

Figure 1 shows a simple depiction of hydrostatic equilibrium.

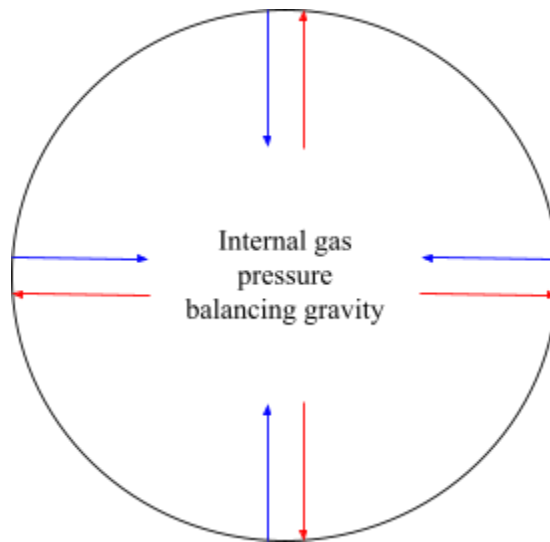


Figure 1: A diagram of gravity and internal gas pressure working against each other equally to satisfy hydrostatic equilibrium. The blue arrows represent gravitational force, pushing inwards, and the red arrows represent internal gas pressure, pushing outwards.

If a star is not in hydrostatic equilibrium, either the pressure forces inside the star or the force of gravity has become stronger than the other, which takes the entire system out of balance. Making sure hydrostatic equilibrium is maintained in the star I will be using for my simulations is important for that very reason.

2.2.2 Astrophysical Timescales

There are three main timescales used in astrophysics, in order from shortest to longest: the free-fall or dynamical timescale, the thermal or Kelvin-Helmholtz timescale, and the nuclear timescale. The free-fall timescale is the time it takes for an object, like a star, to collapse under its own gravity if there is no hydrostatic equilibrium or other forces combating the collapse. The equation for this first timescale is

$$t_{ff} \approx \frac{1}{\sqrt{G\rho}}, \quad (2.5)$$

where ρ is the average density of the object. The thermal timescale is the time it would take for a star to radiate away its total kinetic energy at its current luminosity rate, in other words, how long it would take for a star to cool and contract if nuclear fusion suddenly stopped. The equation for this second timescale is

$$t_{th} \approx \frac{GM^2}{2RL}, \quad (2.6)$$

where M is the mass of the star, R is its radius, and L is its luminosity. The nuclear timescale is the estimate of a star's lifetime on the main sequence based on its energy output via nuclear fusion. The equation for this last timescale is

$$t_{nuc} \approx \frac{\text{fuel supply}}{\text{rate of fuel consumption}} * \text{fraction of star over which fuel is burned}. \quad (2.7)$$

For some sense of comparison, for the Sun, t_{ff} is approximately thirty minutes, t_{th} is approximately 10^7 years, and t_{nuc} is approximately 10^9 years.

2.2.3 Stellar Evolutionary Stages

The evolution of a star is most dependent on its mass, due to the fact that its mass determines how much energy it can produce and how quickly it will do so. The more massive a

star, the shorter its life. There are generally two tracks a star can follow, diverging when its mass is approximately above or below $8M_{\odot}$. Since the star I am using has a mass of $1M_{\odot}$, I will be discussing the track a star with a mass less than $8M_{\odot}$ is expected to follow first.

A star is born when a dense region of gas and dust in space collapses under its own gravitational force. When gravity pulls all of those particles together into a ball of gas, called a protostar, its temperature begins to rise. The denser the protostar becomes, the more heat it traps and generates, becoming hotter and hotter as time progresses. Once a protostar is hot enough to begin turning hydrogen into helium in its core through nuclear fusion, it becomes a MS star. The MS is the stage where a star spends most of its life, as hydrogen is the most abundant element in both the universe and inside stars. A star's outer layers are made up of mostly hydrogen as well. Once a star runs out of hydrogen, it is no longer characterized as a MS star. Figure 2 displays a Hertzsprung-Russell (HR) diagram, a common and convenient way of plotting a star's present stage and mass, and an aid in understanding where a star is most likely to go during its lifetime. In stars $\leq 8M_{\odot}$, conditions in their cores only allow up to helium burning. Once helium burning begins, the star moves onto the giant branch of the HR diagram. The star will expand, becoming cooler on its surface but hotter in its core, and more luminous. Once all the helium fuel is used up and there is no longer enough mass to squeeze carbon to the point of burning, what is left behind is a stellar corpse composed entirely of carbon nuclei so densely packed together that only quantum mechanical effects, electron degeneracy pressure, support the star. These stellar corpses are called white dwarfs, and will inevitably fade over tens to hundreds of billions of years. This is the fate that our Sun will eventually meet.

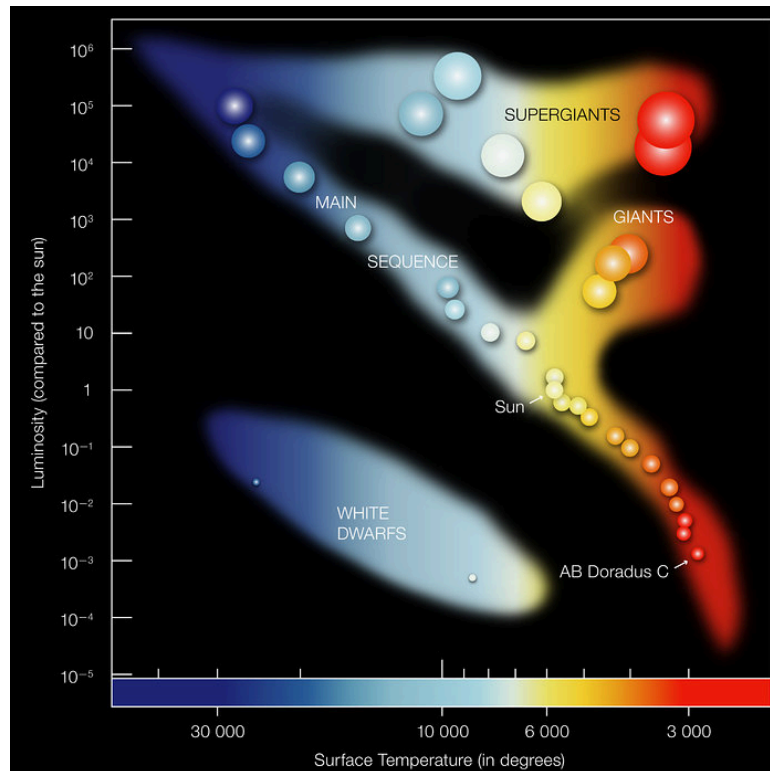


Figure 2: A Hertzsprung-Russell (HR) diagram, which depicts the temperatures of stars plotted against their luminosities. The position of a star in the diagram provides information about its present stage and its mass. Stars that burn hydrogen into helium lie on the main sequence, the diagonal branch in the middle of the diagram [3].

Stars with masses $\geq 8M_{\odot}$ will go through a similar process as mentioned, although much quicker than a sun-like star. These stars do not stop nuclear fusion at helium burning: they have enough mass to burn elements all the way up to iron. The elements inside more massive stars fuse in a fairly standard order: hydrogen $\rightarrow$ helium $\rightarrow$ carbon $\rightarrow$ oxygen $\rightarrow$ neon $\rightarrow$ magnesium $\rightarrow$ silicon $\rightarrow$ iron. Iron is the stopping point of nuclear fusion in stars, as it takes more energy to fuse iron into something heavier than the energy that would come out of the reaction. With nothing left to combat gravity, the star will collapse in an interstellar fireworks show called a supernova. The star essentially implodes, expelling gas, energy, and particles outwards. One of the many beautiful things about our universe is that the gaseous remnants of a star which has

gone supernova will eventually turn into stellar nurseries, creating new stars out of the deaths of others. It is also theorized that all heavy metals were created in the cores of massive stars, and spread across the universe through supernovae, so we are partially, literally, made of stardust. On the lower end of the high-mass spectrum, the stars become neutron stars after going supernova, another version of a stellar corpse supported by electron degeneracy pressure. On the higher end of the high-mass spectrum, gravity overwhelms even quantum mechanical effects, leading to the creation of BHs. Figure 3 below displays a simplified diagram of these two processes.

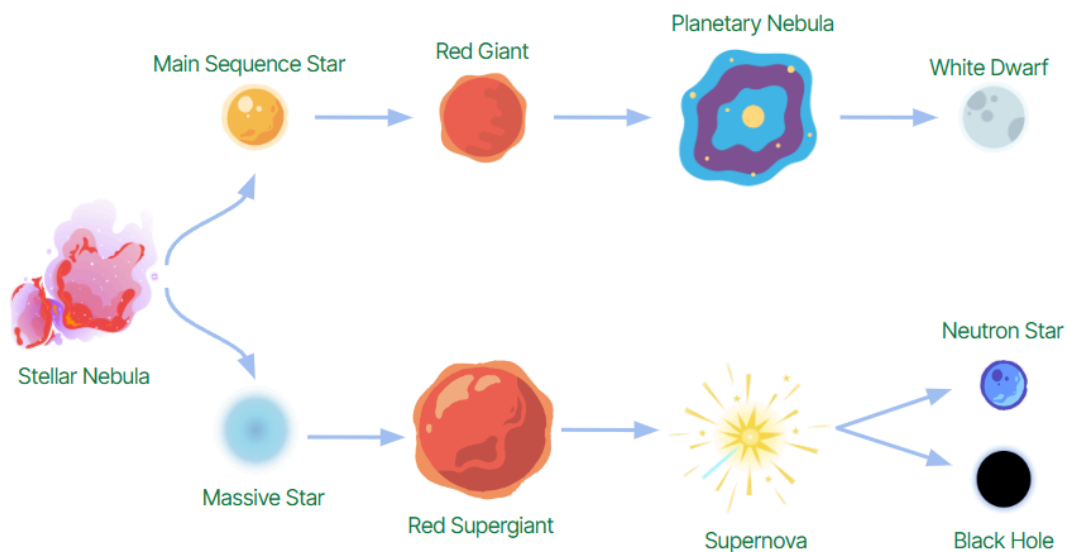


Figure 3: A simplified diagram of the typical stellar evolution process. The top branch is the evolutionary track of a star $\leq 8M_{\odot}$, and the bottom branch is the evolutionary track of a star $\geq 8M_{\odot}$.

In the bottom branch, the outcome of a supernova is dependent on the mass of the star undergoing said supernova. I made this diagram, but the images are from a template found on SlidesCarnival [here](#).

2.3 Black Holes

While we know the process of how BHs are created, they are still incredibly elusive to even the most experienced astrophysicists. BHs cannot be seen with the naked eye, do not produce electromagnetic radiation, and our current understanding of them states they contain a singularity—a point of zero volume yet infinite density—at their centers, with a gravitational pull so strong that nothing can escape them. BHs have only three properties that we can observe: spin, mass, and charge, the same intrinsic properties a particle contains. There are three categories of BHs: stellar mass, intermediate mass (IMBHs), and supermassive (SMBHs). Stellar mass BHs are the most common in our universe, and form directly as a consequence of the collapse of a massive star outlined in Section 2.2.3. Their mass can range from $3M_{\odot}$ to $100M_{\odot}$. There is little direct evidence of the existence of IMBHs at this point, but astrophysicists are fairly confident they do exist. They form most likely through smaller black holes merging with each other, and have a mass range of $100M_{\odot}$ to $1000M_{\odot}$. SMBHs are often found at the center of galaxies, like our very own, and likely form either through galaxies merging with each other or as an extension of the formation process for IMBHs. Their mass range is deemed to be $10^4M_{\odot}$ to $10^9M_{\odot}$ from current observations, but some astrophysicists simplify the range to any BH bigger than $10^4M_{\odot}$. BHs can also exist in binary systems, either with each other or with a star.

2.3.1 Schwarzschild Radius

Once a star collapses within a boundary called the Schwarzschild radius, it becomes a BH. At the Schwarzschild radius, the BH is enclosed by a mathematical surface called the event horizon. Anything inside the event horizon, like the singularity at the BH's center, is hidden from an observer, and once an object passes the event horizon it can never escape the BH's

gravitational force. Additionally, light is unable to escape, meaning at the Schwarzschild radius the escape velocity is equal to the speed of light, $c = 2.998 * 10^8 \frac{m}{s}$. The equation for the Schwarzschild radius of any BH is given by

$$R_s = \frac{2GM_{bh}}{c^2}, \quad (2.8)$$

where M_{bh} is the mass of the BH and c is the speed of light, as mentioned. While we cannot directly observe what happens beyond a BH's event horizon, we can still work out some of their properties and modes of operation.

2.3.2 Tidal Forces and Tidal Radius

The most familiar example of tidal forces are the interactions between the Earth and the Moon that cause ocean tides to rise and fall. The ocean tides depend on the position of the Moon relative to the Earth: when the Moon is on a particular side of the Earth, its gravitational pull (Equation 2.3) causes the Earth to bulge outwards slightly on either side of the planet, making it more of an oblate spheroid shape rather than a sphere. As the Moon orbits the Earth, the bulge in the Earth moves correspondingly to the Moon's position. Since one side of the moon is closer to the Earth than the opposite side, Earth's gravitational force on a small test mass must be greatest on the moon's near side. According to Newton's third law (Equation 2.2), the same situation must also apply to the far side of the Earth because of the moon's gravitational force: this differential force on the Earth is known as a tidal force. This same concept occurs in interactions between stars and BHs. Once a star gets near enough to a BH, the BH's gravity will pull on the star, causing an outward bulge in its hydrogen-dominated outer layers.

The strength of a tidal encounter is characterized by its "penetration factor",

$$\beta \equiv \frac{r_t}{r_p}, \quad (2.9)$$

where r_p is the pericenter distance, or the distance of closest approach between the star and BH, and r_t is the BH's tidal disruption radius. To derive the tidal disruption radius equation, we first need to consider forces of gravity involved in the scenario. The force of gravity between a small piece of mass on the star's surface and the BH's mass is

$$F_{g1} = \frac{GM_{bh}m}{r^2} \quad (2.10)$$

where M_{bh} is the mass of the BH, m is the piece of mass on the surface of the star, and r is the distance between the center points of the star and the BH. We then need to consider the gravitational force between the BH and the star, starting at the surface of the star and traveling into the core,

$$F_{g2} = \frac{GM_{bh}m}{(r-R)^2} \quad (2.11)$$

where all the units are the same as in Equation 2.10, but r is the distance between the surfaces of the BH and the star, and R is the star's radius. By calculating the difference between Equation 2.7 and 2.8 we can approximate the tidal force equation,

$$F_t = 2G \frac{RM_{bh}m}{r^3}. \quad (2.12)$$

The tidal radius is the point at which a star begins to bulge, or be deformed, by the tidal forces. At exactly the tidal radius, we can infer the force of gravity between the star and BH, as well as the tidal force, are equal to one another. Setting them equal to each other, we get

$$F_g = F_t \Rightarrow \frac{GM_*}{R_*^2} = 2G \frac{R_*M_{bh}m}{r_t^3} \quad (2.13)$$

where M_* and R_* are the stellar mass and radius, respectively. Then solving for r_t^3 ,

$$r_t^3 = 2R_*^3 \frac{M_{bh}}{M_*}. \quad (2.14)$$

We can take the cube root of Equation 2.14 and neglect the factor of 2 to obtain

$$r_t = \left(\frac{M_{bh}}{M_*}\right)^{1/3} R_*. \quad (2.15)$$

This equation is an approximation of the tidal radius given the factor of two has been dropped, but it is an accurate enough approximation for my purposes.

2.3.3 Tidal Disruption Events and Partial Tidal Disruption Events

There are two main types of tidal interactions that can occur between a star and BH: a tidal disruption event (TDE), and a partial tidal disruption event (PTDE). A general rule of thumb is that a star will be totally disrupted—the core is ripped apart entirely—if its r_p value falls within one tidal radius, but partially disrupted—the outer layers of the star are stripped away but its core survives—if its r_p value falls between one and two tidal radii. This is not a universal fact, but a helpful approximation for predicting outcomes and making simulations. When PTDEs occur, a star can either be bound or unbound from the system, called a tidal capture or an unbound remnant, respectively.

2.4 Kepler's Laws of Planetary Motion

While Newton's laws of motion and gravitational force lay the basis for classical mechanics, Johannes Kepler's laws of planetary motion lay the groundwork for modern astrophysics. They are summarized as follows: a planet orbits the Sun in an elliptical orbit, with the Sun at one focus of the ellipse; a line connecting a planet to the Sun sweeps out equal areas in equal time; and the average orbital distance of a planet from the Sun is related by the equation

$P^2 = a^3$, where P is the orbital period measured in years, and a is the average distance of the planet from the Sun measured in astronomical units (defined as the average distance between the Earth and the Sun, 1.496×10^{11} m).

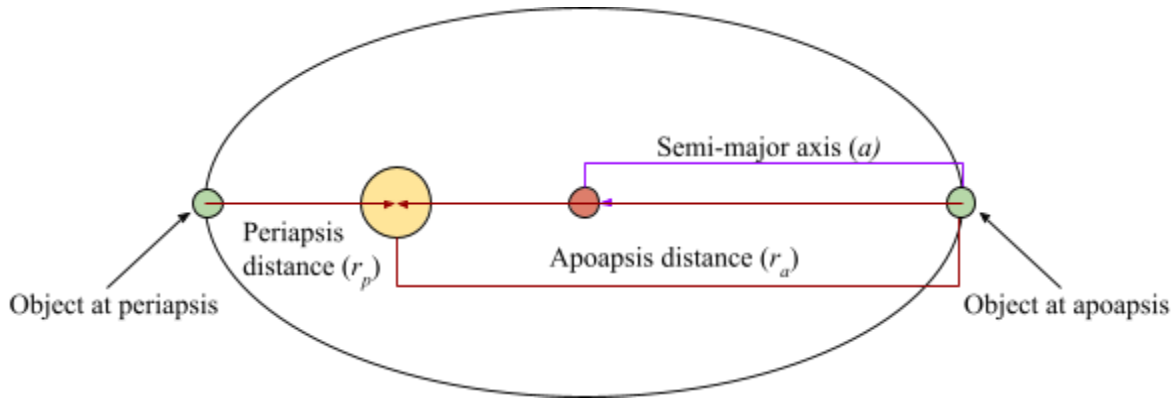


Figure 4: A diagram of an object's elliptical orbit around a second, larger object, with terms labeled. The yellow circle represents the object being orbited, and the red circle is the center of the system.

Kepler's first law turns out to be true for every orbiting body in the universe, and is not exclusive to our solar system: all bound orbits are elliptical. A few important terms to consider when discussing elliptical orbits are periapsis, apoapsis, the semi-major axis, and eccentricity. The semi-major axis (a) is the radius of the orbit, periapsis is the point of closest approach between the orbiting body and focus, apoapsis is the point of furthest approach between the orbiting body and focus, and eccentricity (e) determines how elliptical the orbit is. Figure 4 shows all of these terms on a diagram of an ellipse for ease of understanding, aside from eccentricity. Eccentricity is a value between zero and one, where zero corresponds to a perfectly circular orbit, and one corresponds to a perfectly parabolic orbit. If a system's eccentricity is greater than one, its orbit is hyperbolic, meaning the orbiting body is not bound to the system.

The equation to calculate an object's periapsis distance is given by

$$r_p = a(1 - e), \quad (2.16)$$

and the equation to calculate its apoapsis distance is very similar,

$$r_a = a(1 + e). \quad (2.17)$$

Newton later generalized and derived Kepler's work from his gravitational force law.

Newton's version of Kepler's third law applies to every elliptical orbit in the known universe, not just to our own solar system,

$$P^2 = \frac{4\pi^2}{G(m_1+m_2)}a^3, \quad (2.18)$$

where m_1 and m_2 are the masses of the two objects subject to the elliptical orbit.

3 Methodology

To determine how PTDEs influence stellar evolution in MS stars, and how that evolution may or may not affect further passages the star has around the BH, I will utilize two astrophysical simulation codes: Modules for Experiments in Stellar Astrophysics (MESA), and StarSmasher. MESA will be used to create my star and simulate its evolutionary process, and StarSmasher is what will simulate the interactions between my star and BH.

3.1 Modules for Experiments in Stellar Astrophysics (MESA)

Over the past decade or so, there have been several extremely detailed instruments built and put into use, like the *Parker Solar Probe* and the *James Webb Space Telescope*, which are providing an impressive amount of high-quality data regarding various measurements of stars, strengthening and extending current observational data that can be used in research across the world. Community-driven software like MESA have been transforming how stellar theory and simulations interact with real world observations. MESA, introduced in 2011, is predominantly a stellar evolution code, taking input data describing the characteristics of the star and evolving it

according to modern stellar evolution theory. Different input parameters can be altered depending on the desired star, like its mass, radius, metallicity, etc..

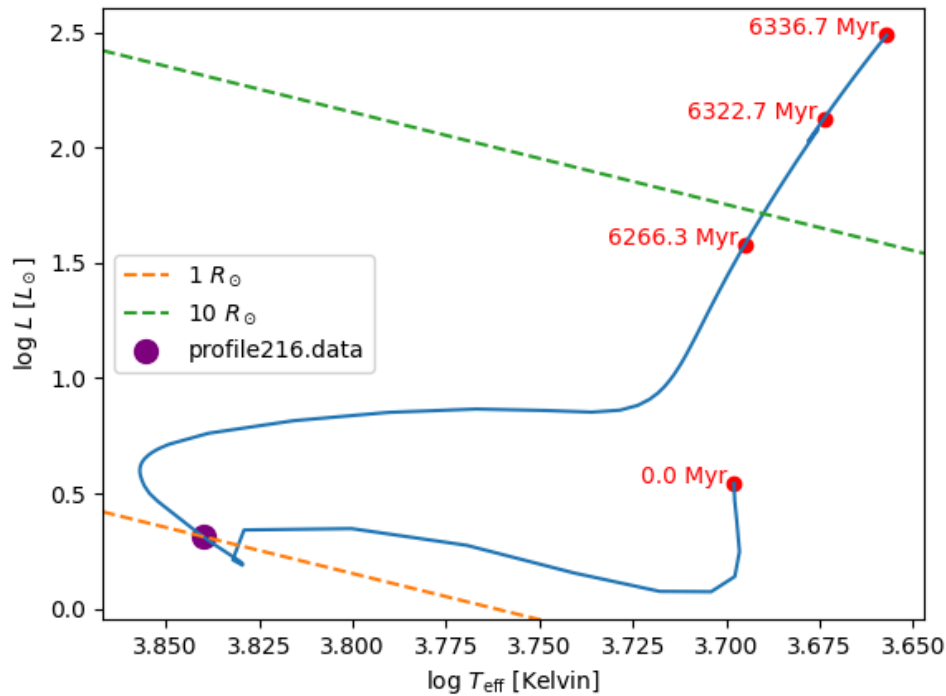


Figure 5: A Hertzsprung-Russell diagram showing the evolution of my $1M_{\odot}$ star. The red dot labeled as 0.0 Myr in the bottom right of the graph is where the star is born, and the red dot labeled as 6336.7 Myr in the top right is where the simulation ended in the giant branch. The purple dot is where the datapoint I selected to use in collisions is located.

After inputting my desired parameters outlined in Section 1.2 into MESA, I let MESA run the code until it created almost 2,000 output files, with each one containing information about the star's current characteristics, like its mass, radius, element content, and more. I looked through the output files to locate one where the star was on the main sequence, and had a radius as close to the Sun as possible, which ended up being profile216.data— its radius was $1.0014R_{\odot}$, but in all future calculations I simplified the value to $1R_{\odot}$. Figure 5 shows the overall evolution of my star in MESA on an HR diagram, and where exactly profile216.data was located on the HR track.

3.2 Smoothed Particle Hydrodynamics (SPH) Code And StarSmasher

The program I will be using to model interactions between my MESA star and BHs is StarSmasher, a Smoothed Particle Hydrodynamics (SPH) code. SPH uses a Lagrangian method to approximate particles as pockets of fluid, where each particle is surrounded by a region called a kernel, which have various parameters like mass, position, velocity, pressure, and temperature. A simple diagram of a kernel can be seen in Figure 6. StarSmasher solves for hydrodynamical forces by calculating the pressure for each particle as a function of the particle's properties, and uses variational equations of motion to calculate the gravitational force between particles. After creating a star in MESA, the desired datapoint can be transferred into StarSmasher to be used in collision simulations.

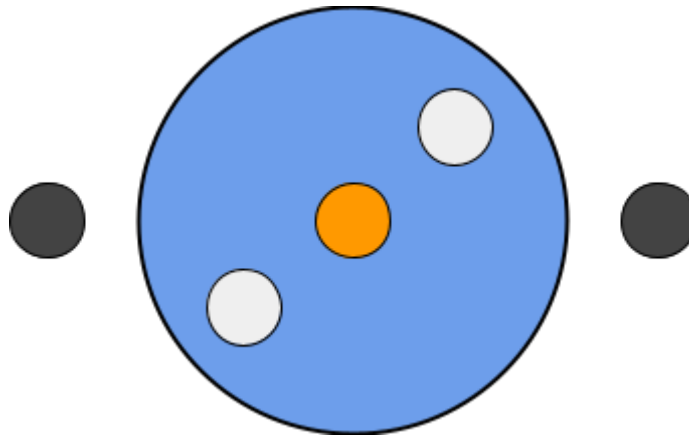


Figure 6: A diagram of a kernel surrounding a particle in StarSmasher. The kernel belonging to the orange particle is represented by the blue circle. Other particles inside the kernel—the white circles (they are called “neighbors”)—are taken into account when calculating properties of the orange particle, but particles outside the kernel—the dark grey circles—have no impact on the properties of the orange particle aside from its gravitational acceleration.

To successfully create collision simulations in StarSmasher, the particles are treated more like fluid-like blobs rather than point particles to create collisions we would more realistically see in observations. StarSmasher reads in given data from the input files to begin running the collisions, and depending on the conditions of the runs, the desired outcome, and how dense the particles are, they can take days to weeks (or even months, in more extreme cases) to finish. The denser the particles become, the slower the output files are generated. In StarSmasher collisions, BHs are treated as point particles with zero spin.

3.2.1 Relaxations

When a MESA star is transferred into StarSmasher, it must be “relaxed” before being put into any collisions. During the transfer, the particles making up the star are left with more energy than is necessary or useful. For the collisions to be realistic, the energy needs to be released, hence the name “relaxation”. At the beginning of a relaxation, the particles are in a rigid lattice. Fully relaxing the star brings the particles to a state of equilibrium, and makes them act more fluidly in a way that StarSmasher can properly account and calculate for. Figure 7 shows the first and last frames of my $1M_{\odot}$ star relaxation.

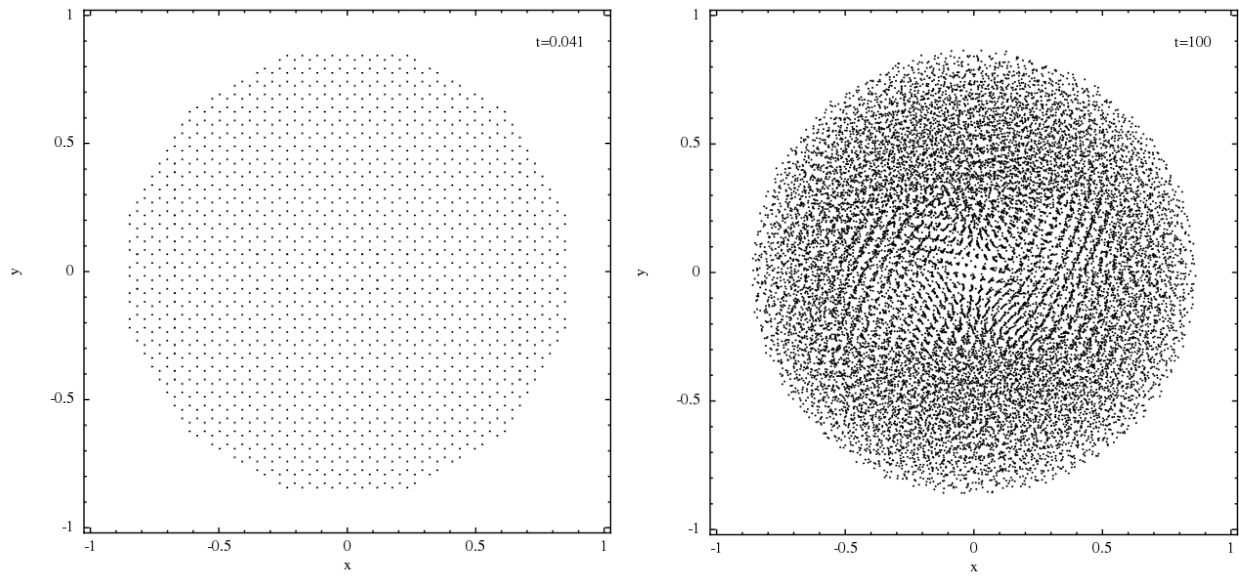


Figure 7: The beginning of my $1M_{\odot}$ star relaxation (left) and the end of the same relaxation (right). It can be seen that the particles on the left are extremely rigid and stiff, while the particles on the right appear much more fluid.

A few different plots can be made in StarSmasher after a relaxation run to check whether it is usable or not, given things can often go wrong if a parameter in the code is incorrect. The most useful are profile plots and energy plots. In profile plots, there are seven separate graphs displaying various properties of the star at the same time. The profile plots for the same $1M_{\odot}$ star as shown in the relaxation can be seen in Figure 8 below.

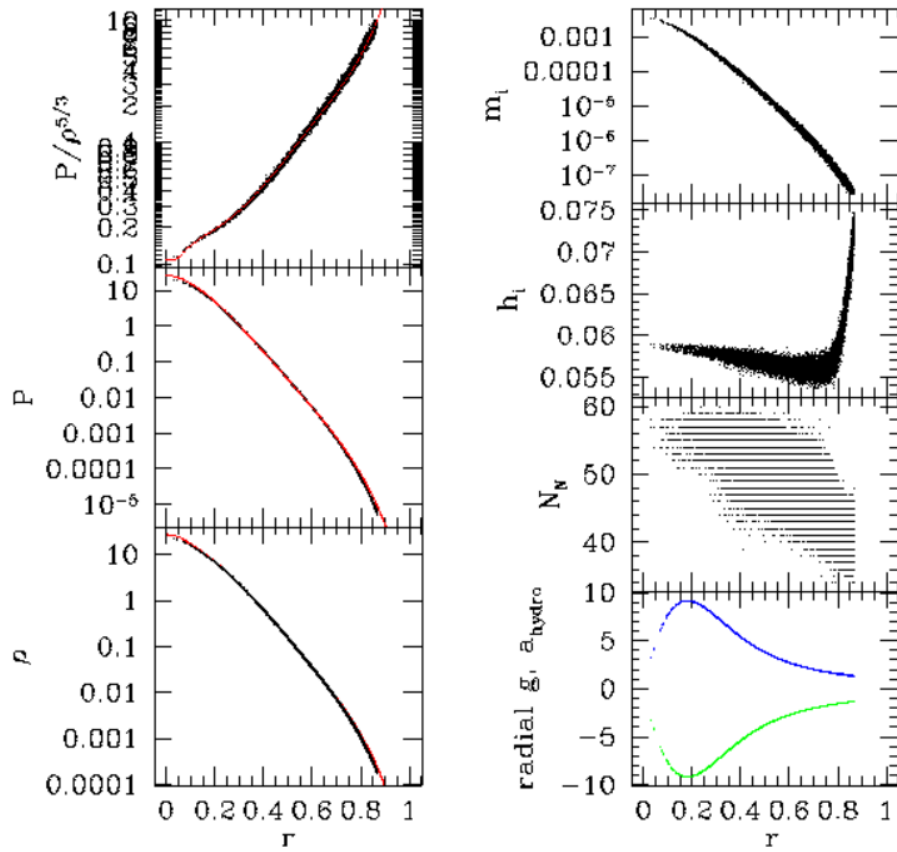


Figure 8: The profile plots of my $1M_{\odot}$ star.

The three graphs on the left of the figure show profiles of pressure (P) and density (ρ), with the very top graph showing the ratio $\frac{P}{\rho^{5/3}}$. The red lines on the left graphs are the lines of best fit, and the black lines are the actual results of the relaxation. The four graphs on the right show properties of the star as a function of its radius. From the top graph to the bottom, m_i is the mass of each individual SPH particle; h_i is something called a smoothing length (half the radius of the particle's kernel, and determined by the number of input particles); N_N is the number of neighbors (how many particles are within two smoothing lengths of any particle); and the green line is radial gravitational acceleration, versus the blue line which is hydrodynamic acceleration.

Ideally, the two lines should mirror each other. The StarSmasher units of density, energy, and time are noted in Appendix A.

Energy plots show the total energy of the star during the course of its relaxation.

Checking to make sure that the energy decreases in the beginning helps confirm whether the star is losing the extra energy it was holding onto after being transferred from MESA to StarSmasher.

It is also important to check that the energy is overall conserved throughout its relaxation run.

The energy plots of my $1M_{\odot}$ star relaxation are shown below in Figure 9.

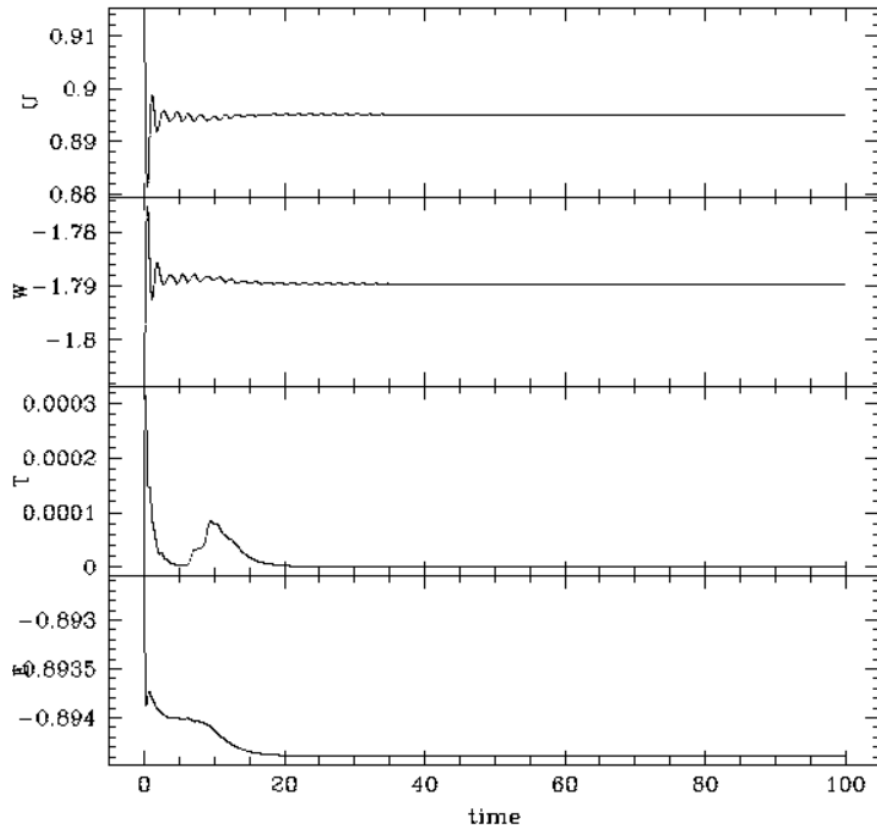


Figure 9: The energy plots of my $1M_{\odot}$ star relaxation. From the top graph to the bottom, U is the internal energy; W is the gravitational potential energy; T is the kinetic energy; and E is the total energy.

4 Collision Results and Analysis

As shown in Section 3.2.1, I made my star in MESA, then transferred it and relaxed it in StarSmasher. Based on the profile plots in Figure 8 and the energy plots in Figure 9, it is a good relaxation. The lines of best fit match up relatively well with the relaxation's actual results, in addition to the fact that the blue and green lines in the radial gravitational acceleration versus hydrodynamic acceleration diagram are well-mirrored. The energy of the star does decrease in the beginning of the relaxation, becomes stable later on, and overall does not change a significant amount, so the energy is essentially conserved. It is not as detailed of a model as it could be, as I used 20,000 particles in the relaxation which is a fairly low amount for StarSmasher, but knowing that the star is normal by astrophysical standards, creating an incredibly detailed model was unnecessary for achieving my goals. I then set up two collision simulations between this relaxed star and a $30M_{\odot}$ BH. The reasoning for choosing a BH mass of $30M_{\odot}$ is that is the median mass of BHs involved in collisions, specifically within dense star clusters.

Collision	Periapsis distance ($R_{\odot}$)	Initial separation ($R_{\odot}$)	Semimajor axis ($R_{\odot}$)
1	3.10723506	20	6687093.205
2	4.66084876	20	6687093.205

Table 1: The initial conditions chosen for my two collision simulations, aside from those already stated like the stellar mass and BH mass. The amount of decimal points in each condition is consistent with what is written in the input file.

The two collisions had all of the same input parameters except for their periapsis distances, where collision 1 had an r_p value equal to r_t and collision 2 had an r_p value equal to $1.5r_t$. The important initial conditions of each collision simulation are summarized in Table 1, and will now be elaborated on. I solved for the black hole's tidal radius using Equation 2.15,

$$r_t = \left(\frac{M_{bh}}{M_*}\right)^{1/3} R_* \Rightarrow \left(\frac{30M_\odot}{1M_\odot}\right)^{1/3} 1R_\odot$$

$$= \left(\frac{30 * 1.989 * 10^{30} kg}{1.989 * 10^{30} kg}\right)^{1/3} 6.9551 * 10^8 m = 2.1611 * 10^9 m.$$

StarSmasher uses solar units for masses and distances, so converting this value to solar radii gives

$$2.1611 * 10^9 m \left(\frac{1R_\odot}{6.9551 * 10^8 m}\right) \approx 3.10723506R_\odot$$

as the tidal radius. Multiplying this value by 1.5 to get the periapsis distance for collision 2 gives a value of $\sim 4.6R_\odot$. The next parameter I considered was the initial separation between the star and the BH, which is simply the distance between the two objects at the beginning of the simulation. I chose to multiply r_t by six to ensure the star and BH have enough distance between them to be non-interacting at first, giving a value of $\sim 18.6R_\odot$, which I rounded up to $20R_\odot$ for simplicity.

The final parameter I considered was how long the orbital period would be. I wanted to ensure that the star would have enough time to make progress in its evolutionary track while on a bound orbit to the black hole. I somewhat arbitrarily chose 1Myr, or 10^6 years, for the orbital period in both collisions, though it would have been preferable to choose an orbital period that is comparable to that star's thermal timescale as discussed in Section 2.3.2: given the star involved in the collisions is very similar to the Sun, an orbital period of roughly 10^7 years would have therefore been more appropriate. Using Newton's version of Kepler's third law, Equation 2.18, I rearranged to solve for the semimajor axis, which I then converted to solar radii and wrote into the collision file.

$$\begin{aligned}
P^2 &= \frac{4\pi^2}{G(m_{bh} + m_*)} a^3 \Rightarrow a = \left(\frac{G^* M_{bh}^* P^2}{4\pi^2} \right)^{1/3} \\
&= \left(\frac{6.67 \cdot 10^{-11} \frac{Nm^2}{kg^2} * 30 M_{\odot} * (10^6 yr)^2}{4\pi^2} \right)^{1/3} = \left(\frac{6.67 \cdot 10^{-11} \frac{Nm^2}{kg^2} * (30 * 1.989 \cdot 10^{30}) * (10^6 * 3.156 \cdot 10^7 s)^2}{4\pi^2} \right)^{1/3} \\
&\approx 4.6479 * 10^{15} m \Rightarrow 4.6479 * 10^{15} m \left(\frac{1 R_{\odot}}{6.9551 \cdot 10^8 m} \right) \approx 6687093.205 R_{\odot}.
\end{aligned}$$

In this calculation, I discarded the m_* term given it was negligible compared to the m_{bh} term.

Images from the two collisions can be seen in Figures 10 and 11 below, in log column density maps.

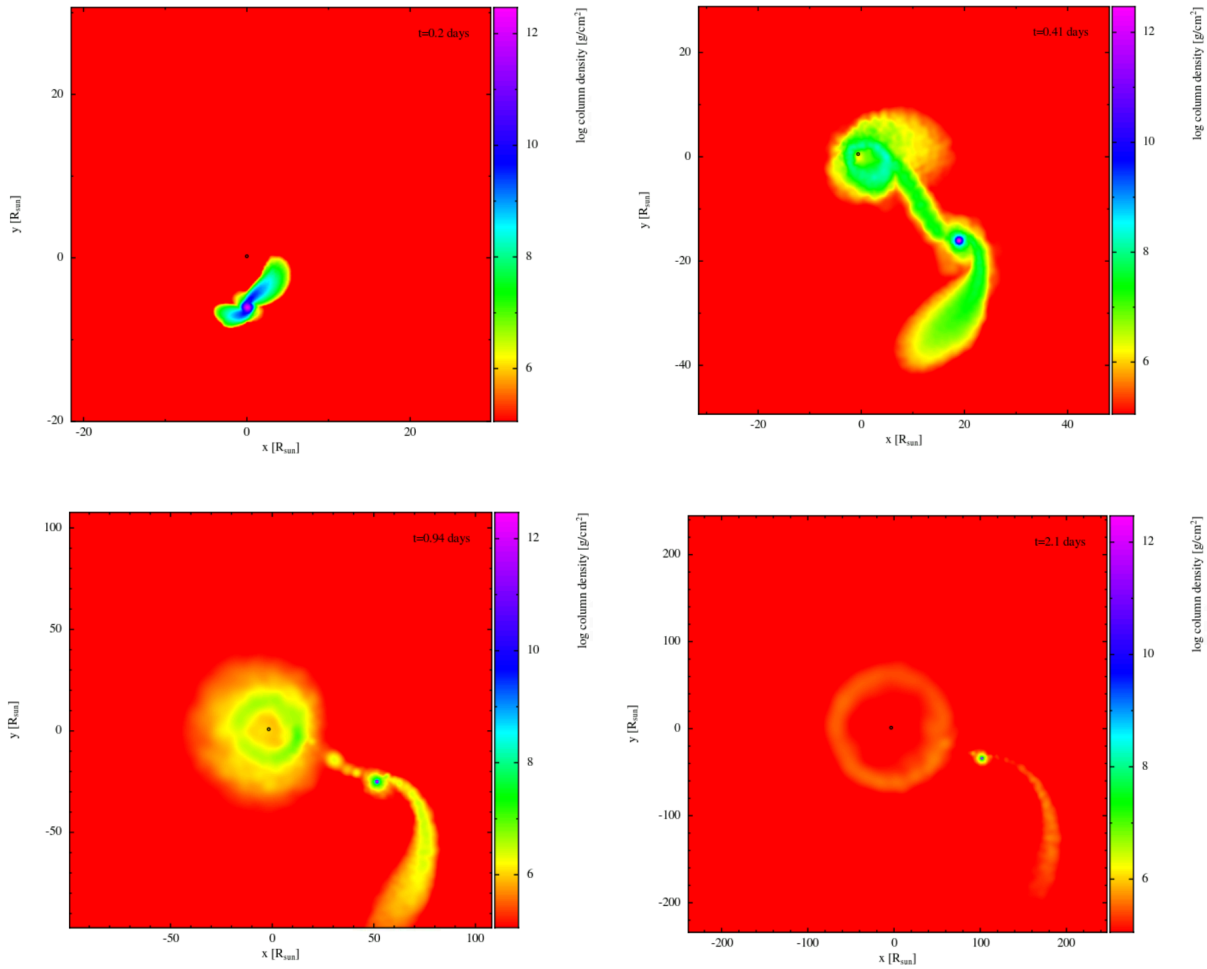


Figure 10: Four snapshot images of collision 1 in chronological order, from the top left panel to the bottom right panel, in a log column density map. The small black dot is the black hole, and the larger, more colorful circle is the star.

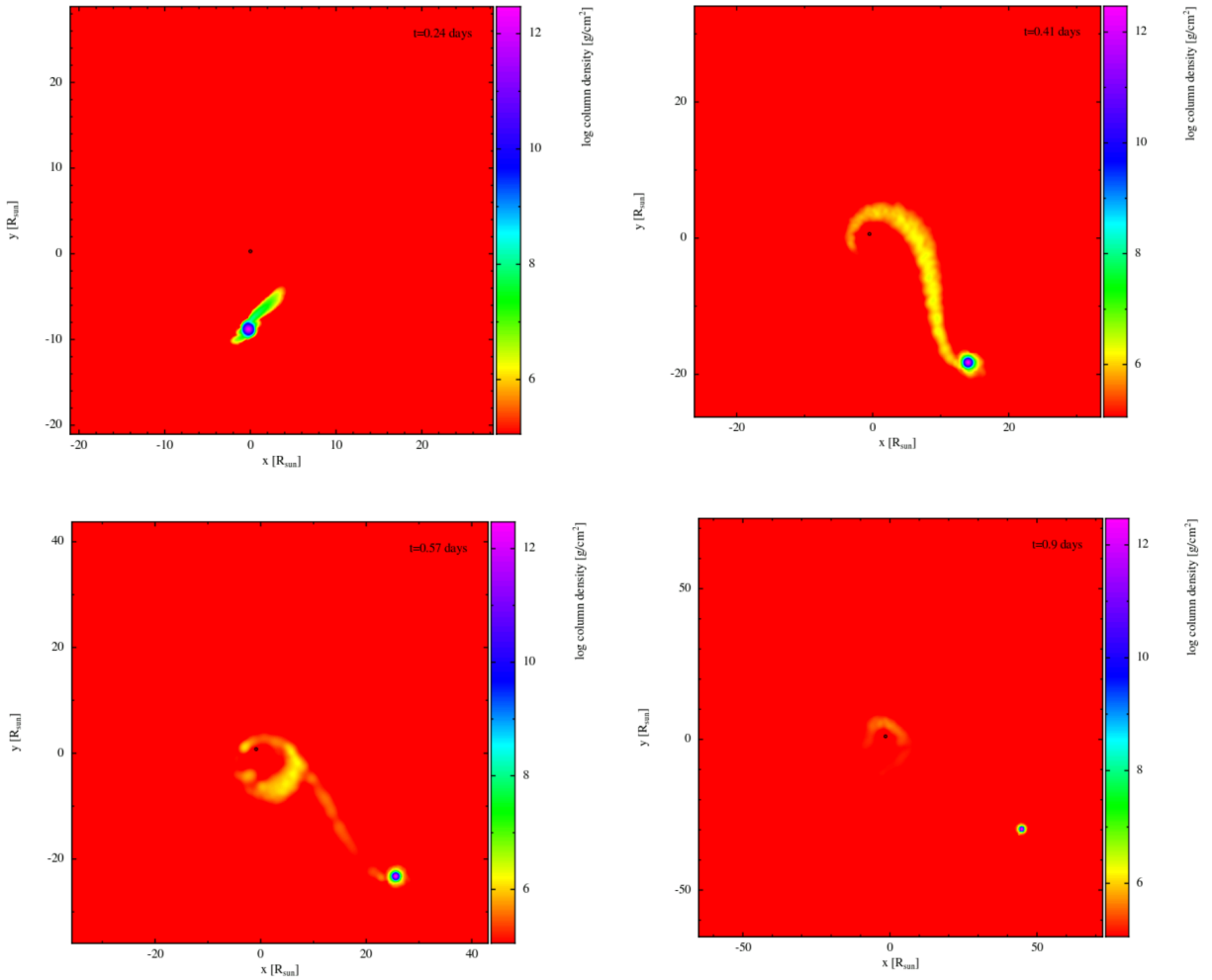


Figure 11: Four snapshot images of collision 2 in chronological order, from the top left panel to the bottom right panel, in a log column density map. The small black dot is the black hole, and the larger, more colorful circle is the star.

I let both collisions run for a long enough time that the inner structure of the star had stabilized, and then went through the process lined out in Appendix B at the end of this paper to transfer both collision products back to MESA. I evolved them for a time as close to their orbital periods as I found possible, which ended up being 1.29Myr rather than a clean 1Myr (it may be possible to edit the MESA code to dump output files at certain time points, though I was unable to figure that out myself). Then, I moved these models back into StarSmasher and re-relaxed them, undergoing the same process as described in Section 3.2.1. I then moved these new relaxed

stars into their respective collision simulations. The simulations that these collision products were put in had the exact same input files as the ones they came from, as stated in Table 1. I was therefore imagining they had been orbiting around the BH for a little over a million years, and then came back around for a second passage with an identical approach to what they had originally. This is very likely not how a secondary PTDE would occur in real life, but with current technological capabilities, this was the simplest way of going about it. Below in Figures 12 and 13 are four snapshots of these secondary collisions.

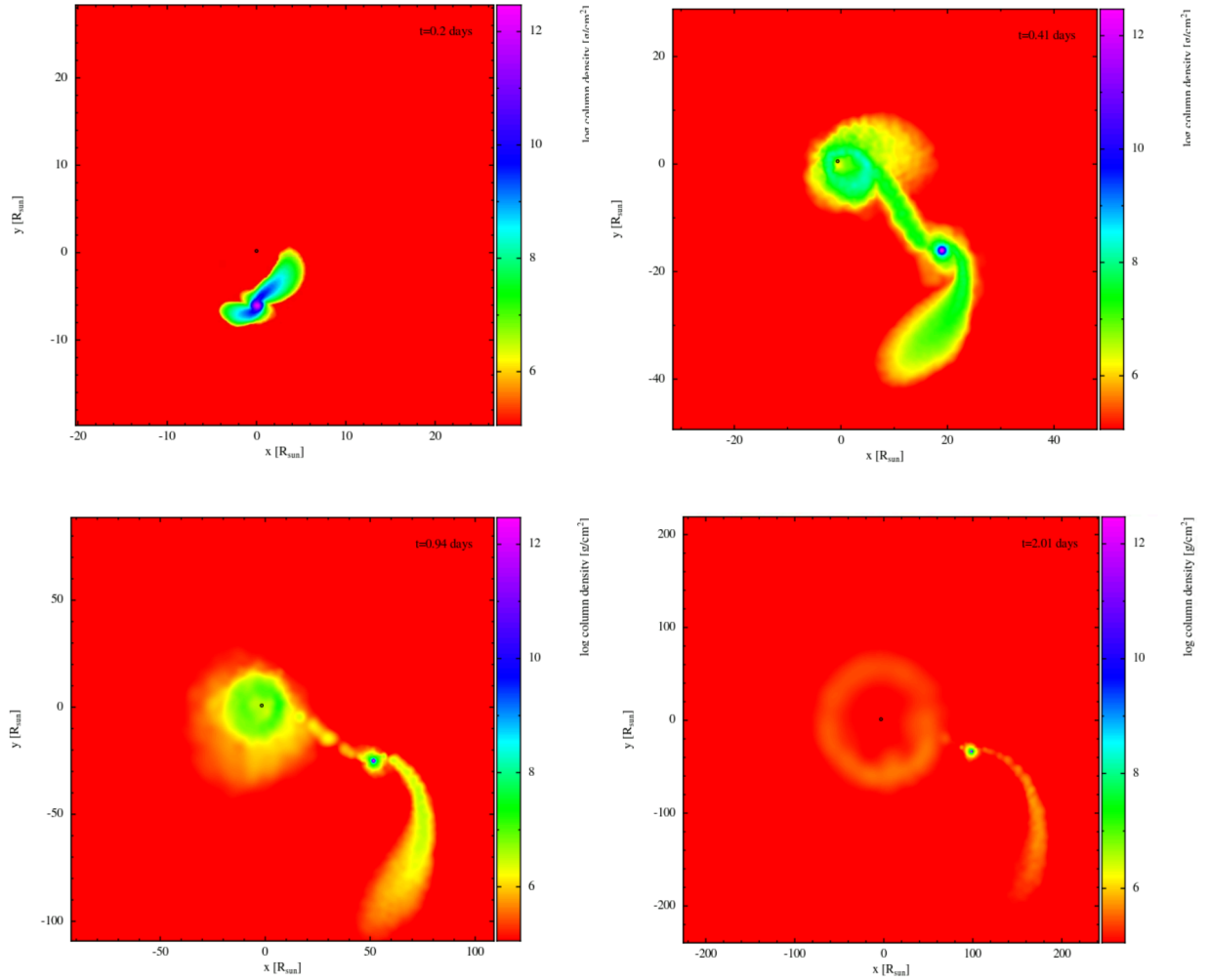


Figure 12: Four snapshot images of collision 1's secondary encounter with the BH in chronological order, from the top left panel to the bottom right panel, in a log column density map.

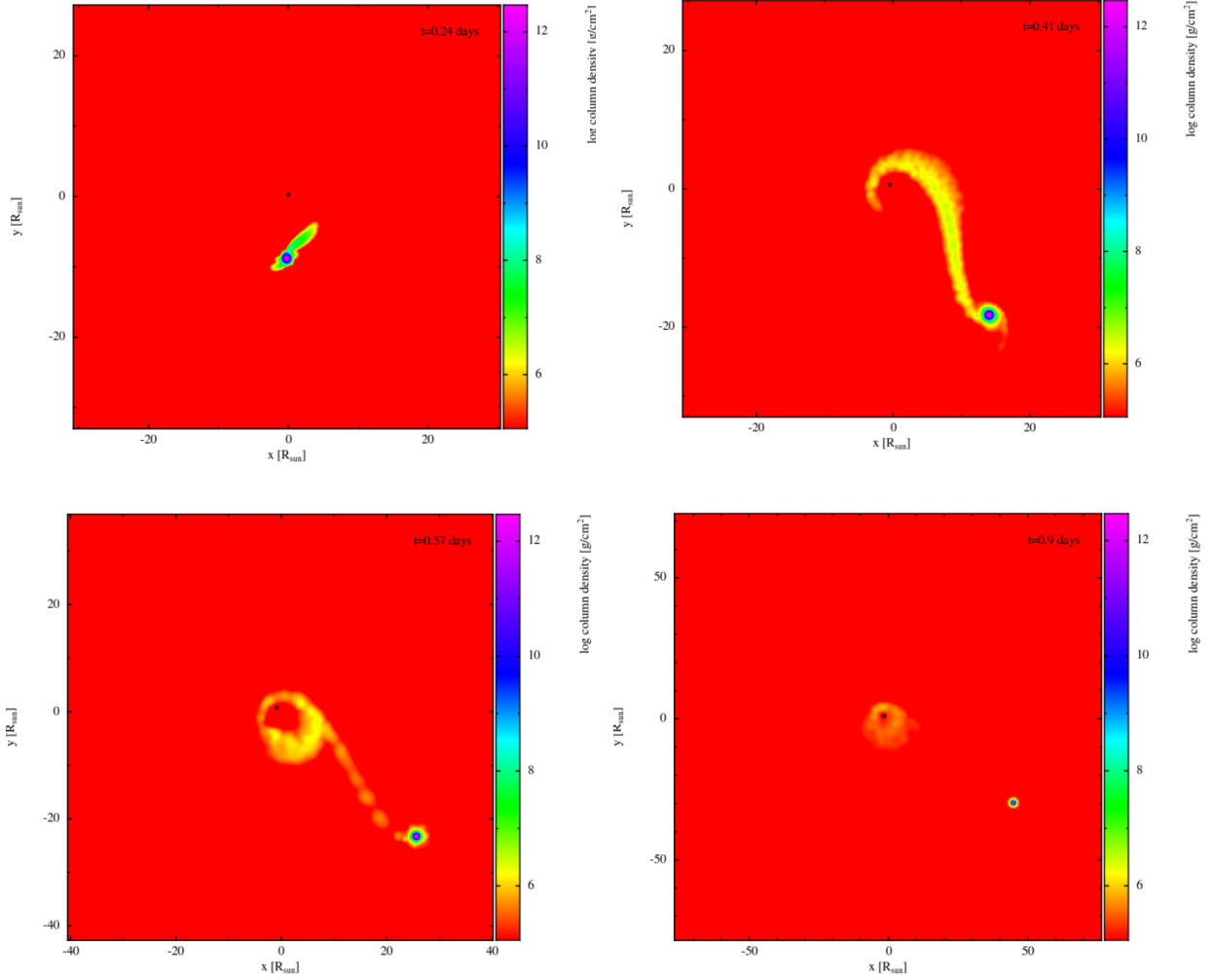


Figure 13: Four snapshot images of collision 2's secondary encounter with the BH in chronological order, from the top left panel to the bottom right panel, in a log column density map.

I then went through the exact same process as written previously: waiting until the internal structures of the stars were stable, going through the process outlined in Appendix B to bring the collision products back into MESA, and allowing them to evolve for another $\sim 1\text{Myr}$ or so. Comparing the first and second passages visually, there is very little difference between them in both collisions. In the images of the second passages, the stars appear to lose a bit more mass than how much they lost from their first passages around the BH, which will be confirmed by

MESA shortly. I then combed through each output file I was concerned with, my original MESA star and both primary and secondary collision products, and tabulated the elemental compositions of each star’s core immediately after the PTDEs and after evolving for $\sim 1\text{Myr}$, which can be seen in Table 2 below. From here throughout the rest of the paper, “CP” will refer to “collision product”.

Property	profile216	CP1- α	CP1- β	CP2- α	CP2- β
Mass ($M_{\odot}$)	1.0000	0.9839	0.9839	0.9999	0.9999
Age (yr)	$2.0949 * 10^9$	$2.0950 * 10^9$	$2.0963 * 10^9$	$2.0950 * 10^9$	$2.0963 * 10^9$
Evolution Time (yr)	$2.0949 * 10^9$	10^5	$1.2916 * 10^6$	10^5	$1.2916 * 10^6$
H1	0.4534	0.4993	0.4991	0.4984	0.4982
He3 ($*10^{-6}$)	6.9205	11.5900	9.7059	10.3211	8.9127
He4	0.5455	0.4996	0.4998	0.5006	0.5008
C12 ($*10^{-6}$)	1.3741	1.112	1.1702	1.1392	1.1989
N14 ($*10^{-4}$)	3.1086	2.8074	2.8075	2.8137	2.8137
O16 ($*10^{-4}$)	3.9945	4.3421	4.3414	4.3346	4.3339
Ne20 ($*10^{-4}$)	1.0425	1.0425	1.0425	1.0425	1.0425

Table 2: The chemical mass fractions of each relevant stellar model’s core from the first passage around the BH compared to the original stellar model, alongside their respective masses, ages, and evolution times. CP1 is in reference to collision 1, where CP1- α refers to the model directly after the collision took place, and CP1- β refers to the model after it evolved for a time roughly equal to its orbital period around the BH. The same naming scheme applies for CP2. The pertinent elements are, in order, Hydrogen, Helium-3, Helium-4, Carbon-12, Nitrogen-14, Oxygen-16, and Neon-20.

There are some changes in both stellar models after the first encounter with the BH compared to the original model, profile216, but there is overall no dramatic nuclear processing

taking place during the million year time frame, just seemingly a redistribution of element abundances. In CP1, the effects from the PTDE likely mixed hydrogen from the star's outer layers into the core, causing the increase of H1 and decrease of He4 after its first passage. The abundance of He-3 increases sharply in both CPs directly after the PTDE took place, possibly from increased nuclear burning due to the star being squeezed from tidal forces, then reduces over time. The amount of C12 and N14 decreases initially in both cases, likely due to some chemical dilution by elemental redistribution throughout the star, though the value of C12 does begin to recover after both stars evolve for a longer amount of time. CP2 displays similar trends to CP1 but the effects are overall weaker, reflecting the less violent PTDE caused by its periapsis distance being further away from the BH than CP1's. There are minimal changes overall in the total mass fractions between both α and β cases, meaning any major compositional changes happened immediately after the PTDEs took place. Subsequent stellar evolution over the ~ 1 Myr time period appears far more gentle and to be governed by normal nuclear processing, rather than additional mixing of the elements or further mass loss. Clearly, PTDEs cause the mixing of elements inside stars, bringing hydrogen into their cores and diluting heavier elements. What differs largely between the two collisions is the amount of mass lost to the BH: CP1 lost significantly more mass than CP2, losing $\sim 1.6\%$, compared to CP2 losing only $\sim 0.01\%$. Where things get more interesting is looking at the results from the secondary BH encounters, which are tabulated below in the same fashion as the previous table.

Property	profile216	CP1- γ	CP1- δ	CP2- γ	CP2- δ
Mass ($M_{\odot}$)	1.0000	0.9658	0.9658	0.9997	0.9997
Age (yr)	$2.0949 * 10^9$	$2.0964 * 10^9$	$2.0977 * 10^9$	$2.0964 * 10^9$	$2.0977 * 10^9$
Evolution Time (yr)	$2.0949 * 10^9$	10^5	$1.2916 * 10^6$	10^5	$1.2916 * 10^6$
H1	0.4534	0.5198	0.5197	0.6448	0.6447
He3 ($*10^{-6}$)	6.9205	12.7083	11.4373	122.2290	78.2751
He4	0.5455	0.4791	0.4793	0.3541	0.3542
C12 ($*10^{-6}$)	1.3741	1.0846	1.1116	23.8984	23.3379
N14 ($*10^{-4}$)	3.1086	2.7472	2.7473	2.2861	2.2926
O16 ($*10^{-4}$)	3.9945	4.4114	4.4109	4.6342	4.6342
Ne20 ($*10^{-4}$)	1.0425	1.0425	1.0425	1.0425	1.0425

Table 3: The chemical mass fractions of each relevant stellar model's core from the second passage around the BH compared to the original stellar model, alongside their respective masses, ages, and evolution times. CP1 is in reference to collision 1, where CP1- γ refers to the model directly after the second collision took place, and CP1- δ refers to the model after it evolved for a time roughly equal to its orbital period around the BH. The same naming scheme applies for CP2. The pertinent elements are, in order, Hydrogen, Helium-3, Helium-4, Carbon-12, Nitrogen-14, Oxygen-16, and Neon-20.

In CP1 after its secondary passage around the BH, there is even more chemical mixing taking place than in the first passage. The amount of H1 in the core continues to increase while the amount of He4 continues to decrease, suggesting a strong influx of H1 into the star's core is occurring. The further increased amount of He3 suggests there is more nuclear processing happening as well. The reduction of C12 and N14 implies that material deeper into the star's core may have been ejected and replaced with fresh material from the outer layers, likely H1. There is minimal compositional changes between CP1- γ to CP1- δ , indicating the core reached a stable structure post-secondary passage after evolving for ~ 1 Myr. What I consider very curious here is

the fact that CP2 was more chemically disrupted than CP1 after their secondary PTDEs, despite the fact that CP1 lost more mass and had an encounter closer to the BH than CP2. This difference is most evident when looking at the spikes of He3 and C12. The large increase of He3 may be caused by a particular He3-dominated shell or localized area in the stellar structure becoming compressed from the tidal interaction. The jump in C12 may be due to unburnt carbon being brought inward towards the core, or a dredge-up of processed nuclear material in the core. From CP2- γ to CP2- δ , the value of C12 remains high and relatively stable, indicating sustained burning or elemental redistribution. This enhanced level of disruption is also clear when looking at the abundances of H1 and He4 in CP2 compared to CP1, displaying how dominant H1 has become over the CP2 core structure. CP1 continued to lose a fair amount of mass from its second PTDE with the BH, losing a total of 3.4%, while CP2 remained largely unbothered, losing only a total of 0.03%. These second passages around the BH amplify the chemical changes in the star's cores, especially in CP2, where more extreme internal changes occurred even with a weaker tidal interaction than in CP1.

Another important thing to note is that the value of Ne20 stayed entirely constant throughout both collisions with the BH. This implies that the densest part of the stars' cores were completely untouched despite the PTDEs occurring. While the chemical compositions of both collision products were shuffled around to some extent, they overall remain consistent with standard MS stars, and if they were able to escape the BH's tidal capture, would continue to evolve relatively normally, as shown in Figure 14 where the HR evolutionary tracks of the original MESA model and all four collision product models are displayed and overlaid on the same diagram.

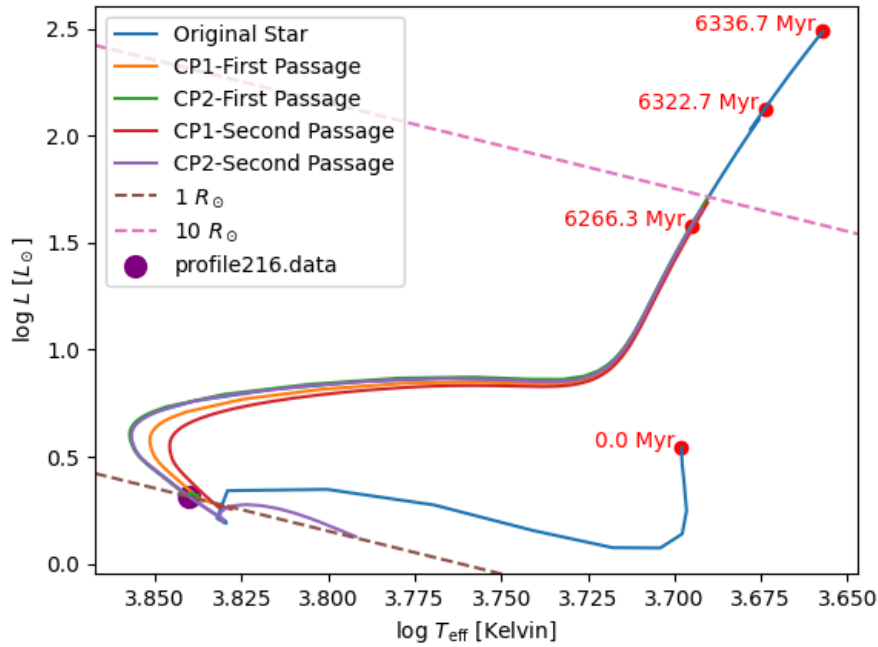


Figure 14: A Hertzsprung-Russell diagram showing the evolution of each relevant stellar model: the original $1M_{\odot}$ star, and the collision products after the first and second passages around the BH. Each track is a different color that is labelled on the diagram. The purple dot is where the datapoint I selected to use in the original collisions is located, the same as in Figure 5.

Looking at Figure 14, it can be seen that both CP2 passages follow the HR track of the original stellar model extremely well from the beginning of their evolutions. Both CP1 tracks follow the same shape as the original HR track, though they are distinctly separate from the original star and both CP2 passages. Visually, they appear to stay on the MS for a slightly shorter amount of time than the other three models, and they additionally lose a small amount of effective temperature. While the two CP1 tracks are distinct at the beginning of their evolutions, by the end of their evolutions close to the $10R_{\odot}$ dashed line, they converge with the other tracks and behave nearly identically. Despite the fact that CP2's second passage model was more chemically disrupted than the other models in regards to its amount of He3 and C12, its evolutionary track is not significantly altered from its first passage model, nor CP1's first or second passages. Clearly, the difference in mass loss from CP1 to CP2 dominates over chemical

composition in terms of how stellar evolution is affected by close encounters with BHs, consistent with what was established in Section 2.2.3.

5 Conclusion

Throughout this research, I analyzed the changes in core chemical abundances in a $1M_{\odot}$ sun-like main sequence star subjected to two partial tidal disruption events by a $30M_{\odot}$ black hole. The data collected corresponds to two particular scenarios: one collision with a periapsis distance equal to the black hole's tidal radius, $\sim 3.1R_{\odot}$ (CP1), and a second collision with a periapsis distance of 1.5 tidal radii, $\sim 4.6R_{\odot}$ (CP2). Each encounter includes snapshots immediately after the first and second passages (α and γ , respectively) and after a subsequent evolutionary period approximately equal to the stars' orbital period around the black hole (β and δ). The reference model, profile216, represents the original stellar core before colliding with anything. Following the initial encounters, the core H1 mass fraction increased in both cases relative to the reference model, while He4 declined correspondingly. This trend indicates that the partially stripped outer hydrogen-rich layers were efficiently mixed into the stellar cores. A modest increase in He3 was also observed immediately after the first encounter in all scenarios, peaking in CP1- α and CP2- α compared to the amount in profile216. These changes likely reflect compression within localized regions of the disrupted star. C12 and N14 decrease slightly in CP1 models, likely due to dilution of the processed core material with less dense material originating from the star's outer layers. In CP2, the changes in these heavier elements are more minute, reflecting the gentler tidal interaction due to its periapsis distance being further away from the black hole than CP1's. Evolution from α to β models over a timescale of $\sim 10^6$ years showed only slight chemical

changes, remaining largely constant overall. This suggests that the majority of restructuring of the core composition occurs immediately following the partial tidal disruption event.

The second tidal encounters induced even more dramatic changes in core composition, especially in CP2. A particularly odd outcome is the sharp rise of He3 in CP2- γ , reaching an order of magnitude higher than the already disrupted CP1- γ model. This could signal extreme localized heating induced by the second passage's compression, or further structural reconfiguration of the core. Notably, CP2 also exhibits a dramatic increase in C12, in contrast with the decline observed in CP1 and primary passage models. This likely reflects that unburnt carbon is being brought inward, or there is a dredge-up of processed nuclear material in the core. The cause of these elemental changes may originate from CP2 losing less mass than CP1, meaning it may potentially have undergone more chemical mixing due to the tidal forces. The C12, N14, and O16 abundances remain stable in both cases, suggesting the stars are experiencing sustained nuclear burning in a rejuvenated core configuration once having the time to settle. While there were some differences inside the stars from the first to second passages, the behavior of the passages themselves were not evidently different from one another.

These results highlight that partial tidal disruption events can significantly modify the internal composition of main sequence stars, though in these two particular cases the partial tidal disruption events did not largely affect the stellar evolutionary processes of those same stars. Even in cases of weaker encounters, i.e. CP2, cumulative effects from multiple passages can produce cores that are rich in hydrogen and heavier elements. The compositions of both collision products were altered, but overall remain consistent with the make-up of standard main sequence stars, and would continue to evolve relatively normally if able to escape the tidal forces of the black hole. Stars surviving multiple, more severe tidal disruptions may contribute to exotic

stellar populations near the supermassive black holes that bind galaxies together, and their mass fraction abundance patterns could provide indirect evidence of tidally-driven stellar evolution.

5.1 Future Work

There are a few different avenues for continued research on this topic. It would be interesting to see how similar encounters would occur between the star and an intermediate mass or supermassive black hole, and how a wider variety of periapsis distances would affect the interactions. Observing the effects on a star undergoing a partial tidal disruption event with a periapsis distance less than one tidal radius would be very interesting: can the star get close enough that its neon core is affected, but still survives after the encounter, or not? Seeing the effects on main sequence stars with different properties, like a larger or smaller mass and/or radius, different metallicity content, etc., would also be worth looking at. There are a ton of input parameters that can be altered in this research, and I unfortunately could not explore all of them, but I believe it would be fun for someone else to.

Like I mentioned early on in Section 4, I somewhat arbitrarily picked 1Myr as the orbital period of the star around the BH by assuming that would be enough time to obtain intriguing results (which I personally believe ended up being true), but an orbital period more comparable to the individual star's thermal timescale would have been more apt. It would be useful to examine the composition of the entire star rather than just the core, and be able to compare the structure of the stellar surface in addition to the structure of the core. I would also like to see if these results can be recreated given how odd CP2's ultimate composition ended up being, mainly in reference to its abundance of He3 and C12.

Appendix A: Constants and StarSmasher Units

Constants	
Solar mass	$1M_{\odot} = 1.9891 * 10^{30}\text{kg}$
Solar radius	$1R_{\odot} = 6.9551 * 10^8\text{m}$
Astronomical unit	$1\text{AU} = 1.496 * 10^{11}\text{m}$
Gravitational constant	$G = 6.673 * 10^{-11} \text{ N}\cdot\text{m}^2\cdot\text{kg}^{-2}$
Speed of light	$c = 2.998 * 10^8 \text{ m}\cdot\text{s}^{-1}$
StarSmasher Units	
Density	$\rho = 1M_{\odot} * 1R_{\odot}^{-3}$
Energy	$E = GM_{\odot}^2 * R_{\odot}^{-1}$
Time	$t = (R_{\odot}^3 * (G^{-1}\cdot M_{\odot}^{-1}))^{1/3}$

Appendix B: Steps to Evolve a Collision Product Between a Star and Black Hole

1. In the directory where your stellar relaxation lives, you will want to evaluate it using a code called `splot` to create an `sph.composition` file that you will use further on— you should be able to access `splot` by copying it from Dr. Lombardi’s directories to your own, if you do not already have it available. Begin by entering `splot` into the linux terminal, which will pull up a menu of options from 0 to 13. Choose option 11, “complete relaxation reconstruction”. `splot` will then ask you for a “starting out file” and an “ending out file”, and for both you will enter the number of last output file from your relaxation. For example, if your relaxation produces up to `out0100.sph`, enter 100 for both the “starting out file” and “ending out file” prompts. When `splot` prompts for your desired step size, enter 1. `splot` will then ask which type of input file you would like to feed to MESA: choose DT, density-temperature. The final element you will be prompted for is “which star are you analyzing?” to which you enter 1, as there is only one star you are concerned with here. In summary, the series of values you will enter into your terminal are

```
splot
11
last sph.out number
last sph.out number
1
DT
1
```

2. You should now have an `sph.composition` file in your relaxation directory. Copy that file over to whatever directory your collision lives in — make sure your collision is using the same relaxed star you made an `sph.composition` file for. Once you copy the file over, make a backup of `sph.composition`, and name it something like `sph.composition_orig` or `sph.composition_backup`: this is not a necessary step, but is a safeguard in case something goes wrong with your `sph.composition` file later. To deal with the black hole in the collision, add a line at the bottom of the `sph.composition` file that consists of eight zeros, each separated by a space. Make sure to hit enter at the end to make a new line, then save and close your `sph.composition` file.

3. Next, you need to remove lines in the `sph.composition` file that correspond to some of the star’s particles that were “eaten” by the black hole. To do this, type

```
grep throwing log*.sph
```

into the linux terminal. This should return multiple particle indices. After the command finishes running, type

```
emacs +(index number) sph.composition
```

where the index number is the first particle index returned by the `grep` command. This will open your `sph.composition` file to that corresponding line: remove it using `ctrl+k`, then backspace, then save the file and close it. Repeat this process for every particle index received.

4. Now run `splot` to determine the components of your star. In the menu selection, choose option 4, “determine components”. When prompted for the “starting out file”, enter 0, and for the “ending out file”, enter the number of the last output file from your collision. When `splot` prompts for your desired step size, enter 1. The program will then ask you how often you want `comp*.sph` files to be dumped: you could enter whatever value makes the most sense for

however many out*.sph files you have, but I always enter 1 here. The final bit of information splot will ask you for is the fraction of internal energy to use in the calculations when determining bound mass. Entering 0 here simplifies the process and makes it so that the code determines boundness based on if there is enough kinetic energy in a particle to overcome the gravitational potential energy. The series of values you will enter into your terminal are therefore

```
splot
4
0
last sph.out number
1
1 (or whatever frequency you personally want comp*.sph
files to be dumped)
0
```

5. Now run splot again to analyze the collision structure of the star. In the menu selection, choose option 3, “generate MESA files - smoothing”. Enter the number of the last output file from your collision for both the “starting out file” and for the “ending out file” like in step 1. When splot prompts for your desired step size, enter 1. Choose DT for the type of input file you would like to feed to MESA, and then enter 1 for “which star are you analyzing?”. The series of values you will enter into your terminal are therefore

```
splot
3
last sph.out number
last sph.out number
```

1

DT

1

6. Run `splot` a final time to generate MESA input files necessary for evolving your collision product. In the menu selection, choose option 10, “write MESA inlist”. Enter the number of the last output file from your collision for both the “starting out file” and for the “ending out file” like in step 1 and 5. When `splot` prompts for your desired step size, enter 1. Choose DT for the type of input file you would like to feed to MESA, and then enter 1 for “which star are you analyzing?”. This will create a directory called `MESA_inlist_files1` which contains a few files, like `entropy.dat` and `angular momentum.dat`. In summary, the series of values you will enter into your terminal are

```
splot
```

```
10
```

```
last sph.out number
```

```
last sph.out number
```

```
1
```

```
DT
```

```
1
```

7. You now have everything you need to evolve your collision product in MESA. On whatever machine you prefer and/or has the most updated version of MESA (which in my case was `carrv122`), make a directory for your MESA run. You want this directory to contain files from a

“raw_mesa_template”: I acquired my raw_mesa_template by copying it from Charlie Gibson’s files to my own, by doing something like

```
cp -pr /data/gibson01/raw_mesa_template/
/path_to_desired_mesa_directory
```

Then copy all the files inside MESA_inlist_files1 to your desired MESA directory. You can do this by typing something in your StarSmasher directory like

```
cd MESA_inlist_files1
scp -pr * carrv122:/path_to_your_mesa_directory
```

8. Run your MESA evolution by entering in your MESA directory terminal

```
initmesa
./clean
./mk
./rn
```

Two recurring issues I had during this process had to do with the star’s entropy once transferring the collision product to MESA and attempting to run the evolution. The first was getting an error message that read

```
stopping because of problems dt < min_timestep_limit.
```

If this error appears, the solution is editing a line in the inlist_project file. There is a line called “timescale_for_entropy_relax” which is set to some value like “1d-1” where the first number is the multiplier, and the “d” stands for 10^x (i.e. 1d-1 equals $1 * 10^{-1}$). This is the timescale on which energy is being injected into the shells of the zone. If you increase the number by some

large factor, then it will converge easier. I always chose to increase the number to 99, i.e. 1d99 instead of 1d-1.

The second was getting an error message that read

```
failed in eosDT_get
do_relax_initial_entropy ierr          -1
do_star_job_controls_after ierr        -1
do_before_evolve_loop ierr            -1
do_before_evolve_loop ierr            -1.
```

This error has a less simple, and more tedious, solution. You need to return to the directory where the collision occurred, and redo steps 5 and 6, picking DE, density-internal energy, rather than DT. Subsequently, you need to delete the original MESA_inlist_files1 file, as well as the original *.dat files. You can then continue on to steps 7 and 8 as normal.

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