Evolution of massive stars

as progenitors of neutron stars and black holes

중력파 여름학교 7. 26. 2022

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광고

서울대학교 천문전공 대학원 입학 설명회

날짜: 2022년 9월 16일 금요일

장소: 서울대학교 19동 (Zoom도 병행할 수 있음)

시간과 구체적 내용은 천문전공 홈페이지를 통해 추후 공지 예정

홈페이지: astro.snu.ac.kr

Evolution of massive stars as neutron star and black hole progenitors

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1. 항성진화의 기본 원리들
2. 무거운 별의 진화 과정
3. 무거운 별의 죽음: 초신성 폭발 혹은 블랙홀 형성의 조건
4. ~~쌍성계의 진화와 중성자별/블랙홀 쌍성계의 형성~~

I. Basics of stellar evolution

Stellar luminosity with black-body approximation

- If you integrate the Planck function over the whole frequency range, you get the Stefan-Boltzmann law as the following:

$$F = \sigma T^4 \quad \text{at stellar surface}$$

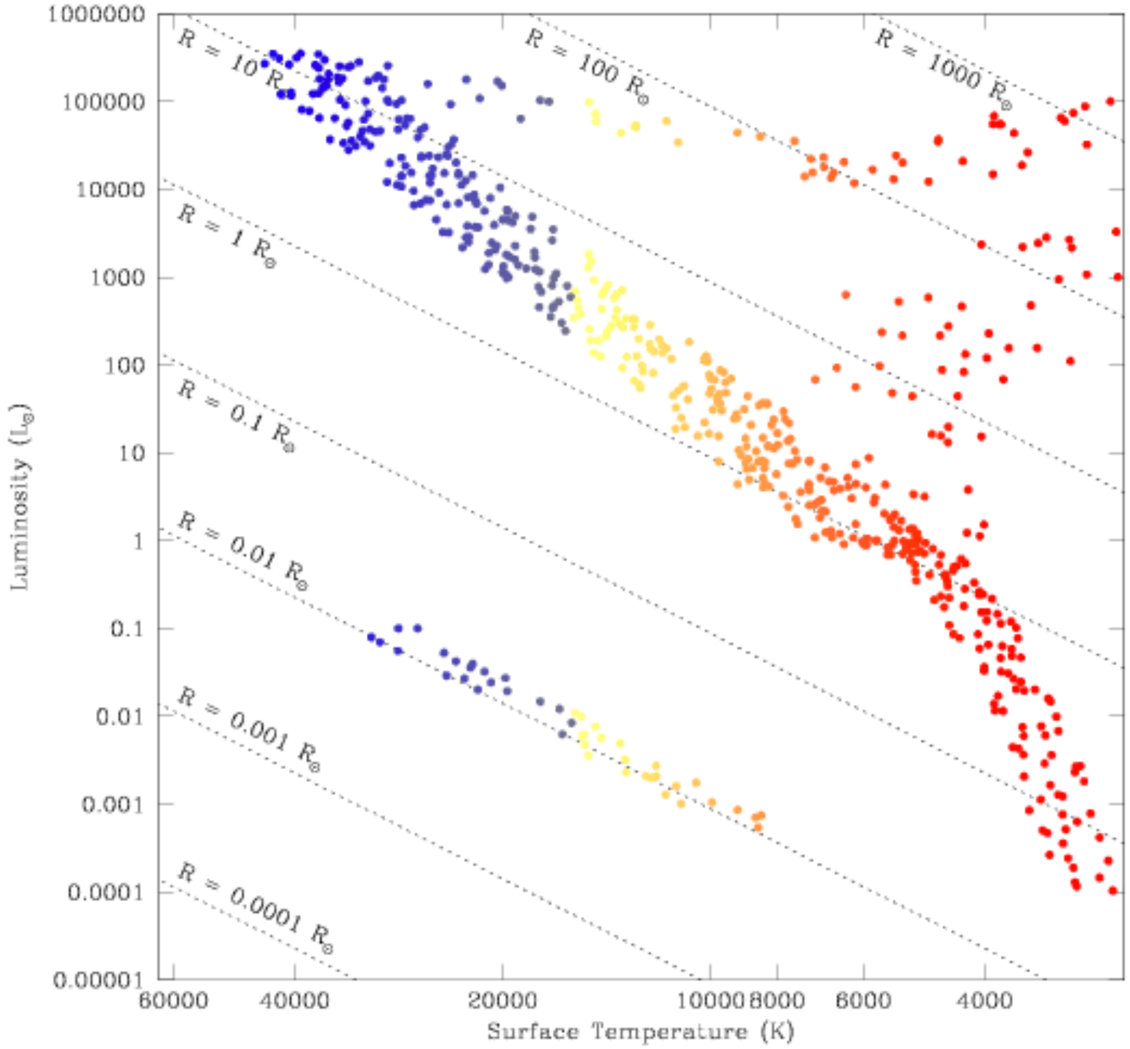
- Here F = flux (energy per unit area per unit time)

- The stellar luminosity (total energy per unit time) is then given by:

$$L = 4\pi R^2 F = 4\pi R^2 \sigma T^4$$

- This means: stars are bright because they are hot (regardless of their energy source).

Hertzsprung-Russell Diagram

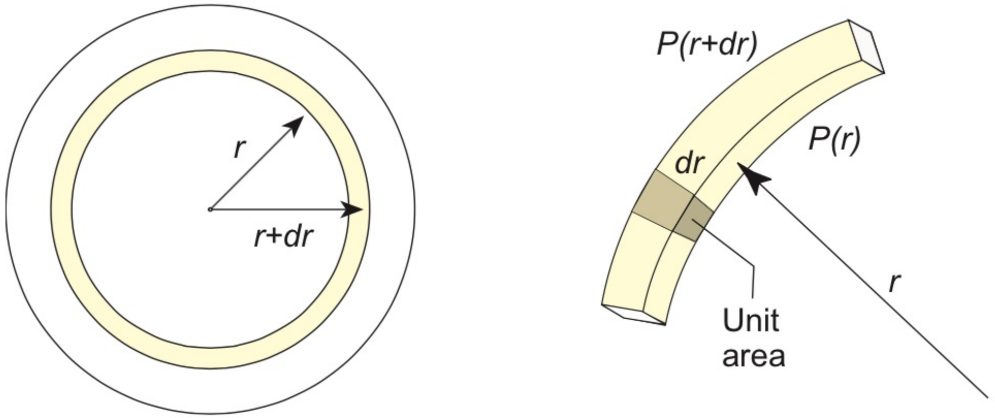


Three principles that govern the evolution of stars

- Stars are in hydrostatic equilibrium
- Stars have a negative heat capacity
- Stars lose energy by radiation

1. Stars are in hydrostatic equilibrium.

Hydrostatic Equilibrium Equation



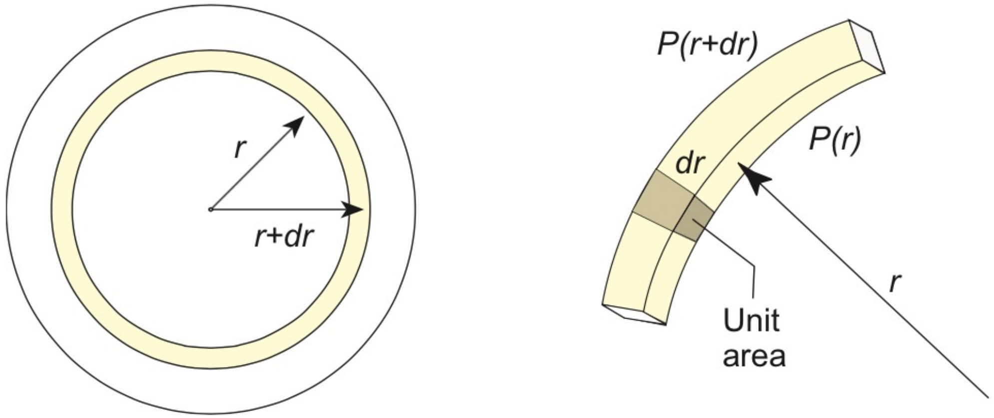
The inward gravitational force acting on the unit mass $dm = \rho dr^3$:

$$F_{\rm Grav} = -\frac{GM_r dm}{r^2} = -\frac{GM_r \rho dr^3}{r^2}$$

The outward pressure force acting on the unit area:

$$F_{\rm pressure} = dr^2[P(r) - P(r+dr)] = -dr^2 \frac{\partial P}{\partial r} dr$$

Hydrostatic Equilibrium Equation



By equating the two: $$F_{\text{grav}} = F_{\text{pressure}},$$

we get the equation that describes the hydrostatic equilibrium in stars:

$$\frac{dP}{dr} = -\frac{GM_r}{r^2}\rho = -g\rho$$

Force per unit volume

Equation of State for a non-degenerate gas

$$P = nkT + \frac{1}{3}aT^4$$

Ideal Gas pressure

Radiation pressure

n = number density

k = Boltzmann constant

a = radiation constant

Central pressure and temperature from the hydrodynamic equilibrium

using $\rho \sim \frac{M}{R^3}$ and $\frac{dP}{dr} \sim \frac{P_c}{R}$, we get

$$P_c \sim \frac{GM^2}{R^4} = 1.1 \times 10^{16} \left(\frac{M}{M_\odot}\right)^2 \left(\frac{R}{R_\odot}\right)^{-4} \ \text{dyn cm}^{-2}$$

and

$$T_c = \frac{\mu_c m_u P_c}{k\rho_c} \sim \frac{G\mu_c m_u M}{kR} = 1.9 \times 10^7 \left(\frac{M}{M_\odot}\right) \left(\frac{R}{R_\odot}\right)^{-1} \left(\frac{\mu_c}{0.85}\right) \ \text{K}$$

2. Stars have a negative heat capacity.

Virial Theorem

VIRIAL :
German, from Latin *vires*, plural, strength, power + German *-ial*; akin to Latin *vis* strength, force, violence (from Merriam-Webster)

Hydrostatic equilibrium equation: $$\frac{dP}{dr} = -\frac{GM_r}{r^2}\rho = -g\rho$$

$$\frac{4}{3}\pi r^3 dP = -\frac{4}{3}\pi r^3 \frac{GM_r}{r^2}\rho dr = -\frac{GM_r}{3r}4\pi r^2 \rho dr = -\frac{GM_r}{3r}dM_r$$

If we integrate the above equation, we get the following relation.

$$2E_{\mathrm{int}} = -E_{\mathrm{grav}}$$: Virial Theorem

where we assumed the ideal gas equation of state.

The Virial Theorem and the ρ-T relation in stars

Virial Theorem: $$2E_{\text{int}} = -E_{\text{grav}}$$

Therefore: $$\frac{3Mk\bar{T}}{\mu m_u} = \alpha \frac{GM^2}{R}$$

And: $$\bar{T} = \frac{\alpha \mu m_u G}{3k}\frac{M}{R} = \frac{\alpha \mu m_u G}{3k}\left(\frac{4\pi}{3}\right)^{1/3} M^{2/3}\bar{\rho}^{1/3}$$

$\bar{T} =$ average temperature of the star

$\alpha =$ some constant that depends on the density profile

$$T \propto M^{2/3}\rho^{1/3}$$

- For a given temperature, density is lower for a more massive star.
- As the star contracts, temperature increases.

Virial Theorem and Heat Capacity of a Star

- Total Energy: $$E_{\mathrm{tot}} = E_{\mathrm{grav}} + E_{\mathrm{int}}$$

- Virial Theorem: $$2E_{\mathrm{int}} = -E_{\mathrm{grav}}$$

$$E_{\mathrm{tot}} = -E_{\mathrm{init}}$$

→ $$\Delta E_{\mathrm{tot}} = -\Delta E_{\mathrm{init}}$$

Stars have a negative heat capacity!

3. Stars lose energy by radiation.

Stars radiate light.

Luminosity in hydrostatic equilibrium ≈ Energy loss rate by radiation. (Energy loss by neutrino emission is not important for usual stars)

$$L = -\frac{dE_{\rm tot}}{dt}$$

If we ignore any internal energy source, from the virial theorem,

$$E_{\rm tot} = -E_{\rm int} = \frac{1}{2}E_{\rm grav}$$

we get:

$$L = -\frac{dE_{\rm tot}}{dt} = \frac{dE_{\rm init}}{dt} = -\frac{1}{2}\frac{dE_{\rm grav}}{dt}$$

This means: **Stars in hydrostatic equilibrium would become hotter and more compact as they lose energy by radiation,** if there is no internal energy source.

Kelvin-Helmholtz Timescale

If a star does not have any internal energy source (i.e., nuclear burning), the star would contract as they lose energy by radiation to maintain hydrostatic equilibrium.

The timescale for this ***thermal contraction (this is NOT dynamical free fall)*** is called "Kelvin-Helmholtz" time scale (or thermal timescale).

$$\tau_{\rm KH} = \frac{E_{\rm tot}}{L} = \frac{GM^2}{2RL}$$

$$\tau_{\rm KH} = 1.5 \times 10^7 [{\rm yr}] \left(\frac{M}{M_\odot}\right)^2 \left(\frac{R_\odot}{R}\right) \left(\frac{L_\odot}{L}\right)$$

KH time is much longer than the dynamical time (i.e., sound crossing time or free fall time)

Stars with nuclear burning

L_{nuc} = energy generation rate by nuclear burning

L = energy loss rate by radiation

The virial theorem tells us:

$$L - L_{\rm nuc} = -\frac{dE_{\rm tot}}{dt} = -\frac{dE_{\rm int}}{dt} - \frac{dE_{\rm grav}}{dt} = \frac{dE_{\rm int}}{dt} = -\frac{1}{2}\frac{dE_{\rm grav}}{dt}$$

Thermal equilibrium : the energy loss by radiation occurs at the same rate with the energy generation by an internal source (e.g., nuclear reactions).

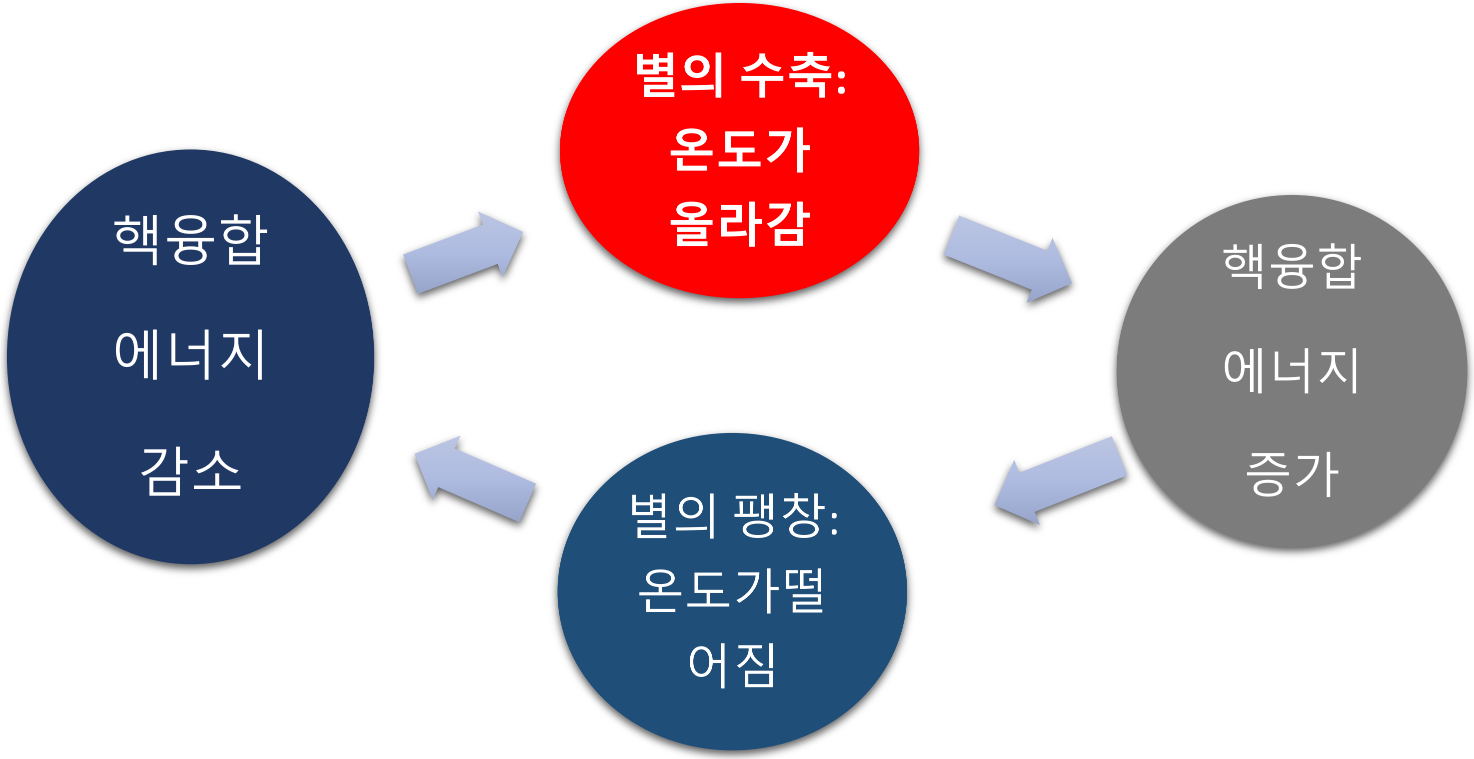
$$L = L_{\rm nuc}$$

in thermal equilibrium for main sequence star.

If there's a perturbation in terms of energy (i.e., instant extra-energy input or loss), thermal equilibrium can be restored on **the thermal timescale**.

Stars with nuclear burning

$$L - L_{\rm nuc} = -\frac{dE_{\rm tot}}{dt} = \frac{dE_{\rm int}}{dt} = -\frac{1}{2}\frac{dE_{\rm grav}}{dt}$$



- $L - L_{\rm nuc} > 0$: the star loses more energy than nuclear energy → contraction
- $L - L_{\rm nuc} < 0$: Stars gain more energy due to nuclear burning than it loses → expansion

Nuclear Burning Timescale

$$\tau_{\rm nuc} = \phi f_{\rm nuc} \frac{Mc^2}{L} \approx 10^{10} \frac{M}{\rm M_\odot} \frac{\rm L_\odot}{L} \ \rm yr$$

$\phi =$ the fraction of the rest mass of the nuclei that is converted into energy
≈ 0.007 for hydrogen burning

$f_{\rm nuc} =$ the fraction of the stellar mass that serve as nuclear fuel

$$\tau_{\rm nuc} >> \tau_{\rm KH} >> \tau_{\rm dyn}$$

Stars can find thermal equilibrium on a very short timescale compared to the nuclear burning timescale.

Energy transport by radiation

$$F_r = -\frac{1}{3} c\lambda \frac{du}{dr}$$

F = photon flux (energy per unit area per unit time)
u = radiation energy density
c = speed of light
λ = mean free path of a photon

$$u = aT^4$$

$$\lambda = \frac{1}{\rho\kappa}$$

$\rho = \text{mass density}$

$\kappa = \text{opacity}$

$$F_r = -\frac{4acT^3}{3\rho\kappa}\frac{dT}{dr}$$

(energy per unit area per unit time)

$$L_r = 4\pi r^2 F_r = -\frac{16\pi acr^2T^3}{3\rho\kappa}\frac{dT}{dr}$$

(energy per unit time)

Dimensional analysis to derive the M-L relation

$$\frac{dP}{dr} = -\rho \frac{GM_r}{r^2}$$

$$L_r = -\frac{16\pi a c r^2 T^3}{3\kappa\rho} \frac{dT}{dr}$$

$$P = nkT = \frac{\rho k T}{\mu m_{\mathrm{u}}}$$

$$\frac{\rho T}{\mu R} \sim \frac{\rho M}{R^2}$$

$$\frac{L}{R^2} \sim \frac{R^3 T^4}{\kappa M R}$$

$$TR \sim M\mu$$

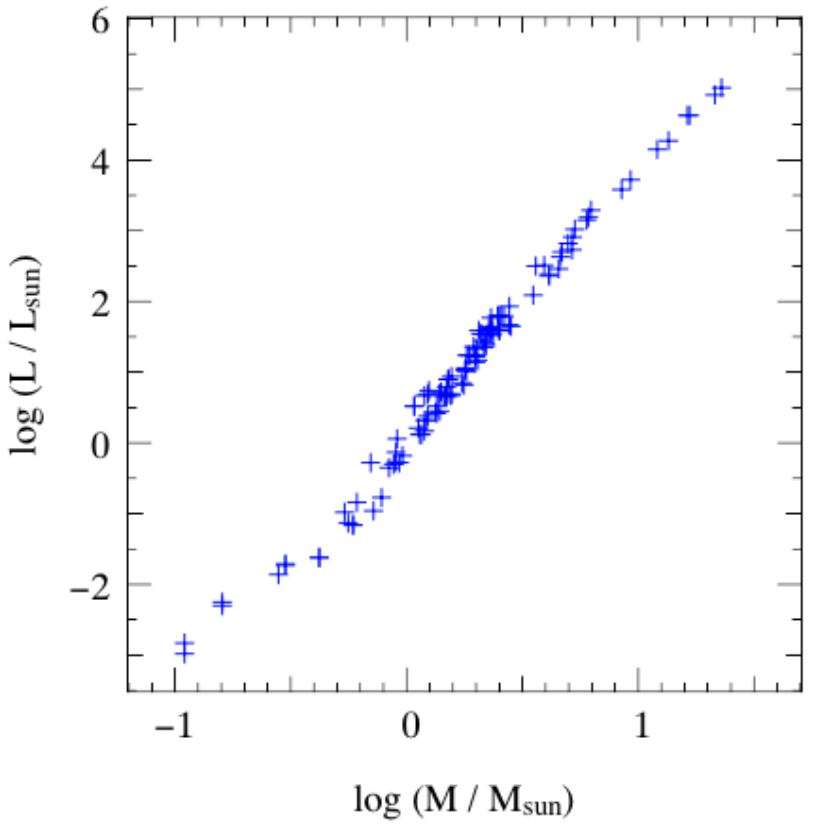
$$L \sim \frac{(RT)^4}{\kappa M}$$

$$L \sim \frac{\mu^4 M^3}{\kappa}$$

for an ideal gas where the gas pressure is dominant.

Mas-Luminosity Relation and Nuclear Burning Time

$$L \propto M^{3.8}$$



$$\tau_{\rm nuc} \propto \frac{M}{L} \propto M^{-2.8}$$

More massive stars have a shorter lifetime.

e.g.

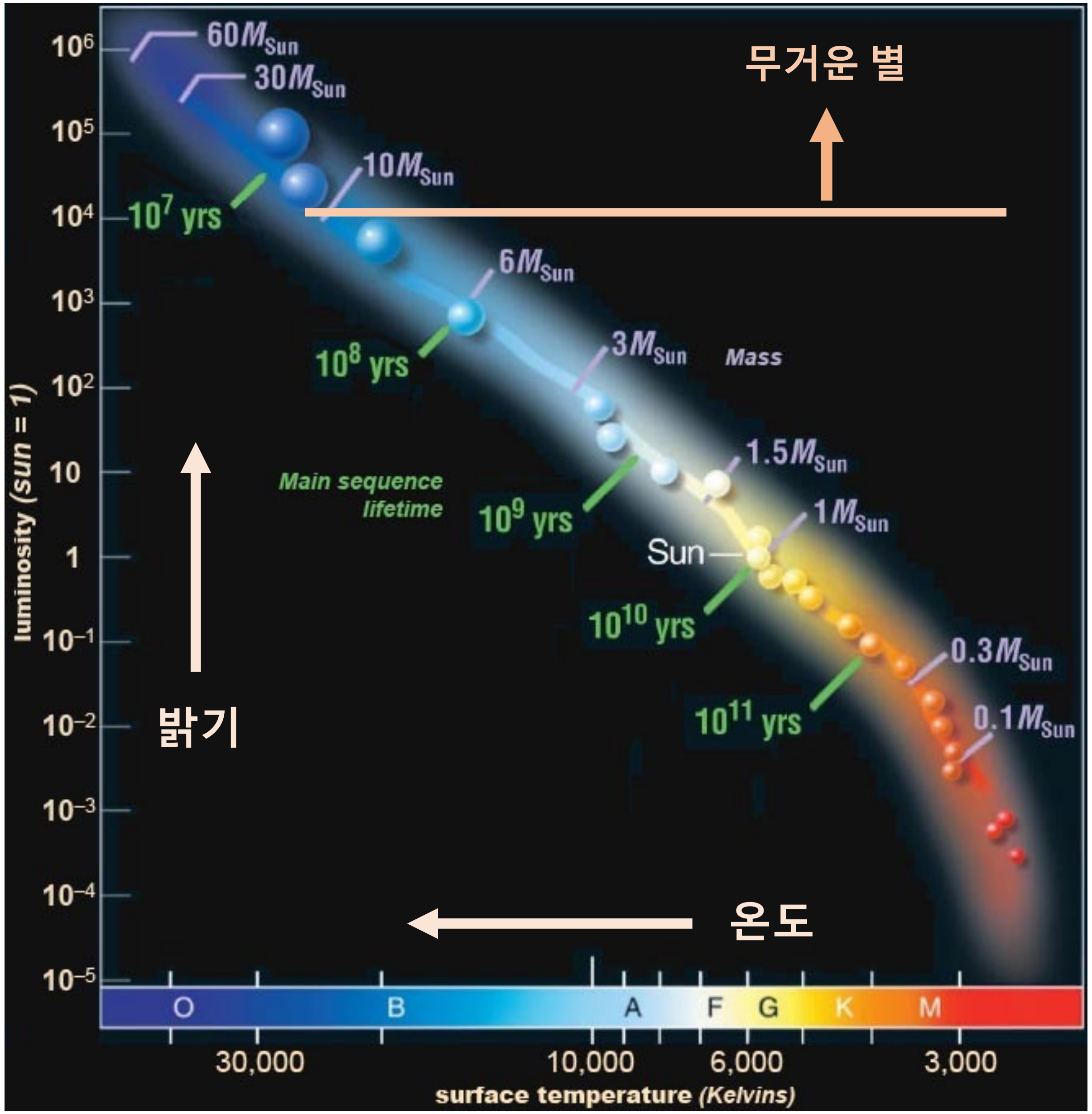
- 1 Msun star : 10 Gyr
- 20 Msun star: 10 Myr

별의 수명

$$수명 = \frac{연료의\ 총량\ (질량에\ 비례)}{에너지\ 손실율\ (밝기)}$$

무거운 별의 수명은 최소 이백만 년에서 최대 삼천만 년 정도

- 태양의 수명: 100억 년
- 지구의 나이: 46억 년
- 태양계 탄생 이후 생명이 등장하는 데까지 필요한 시간: 6억 년

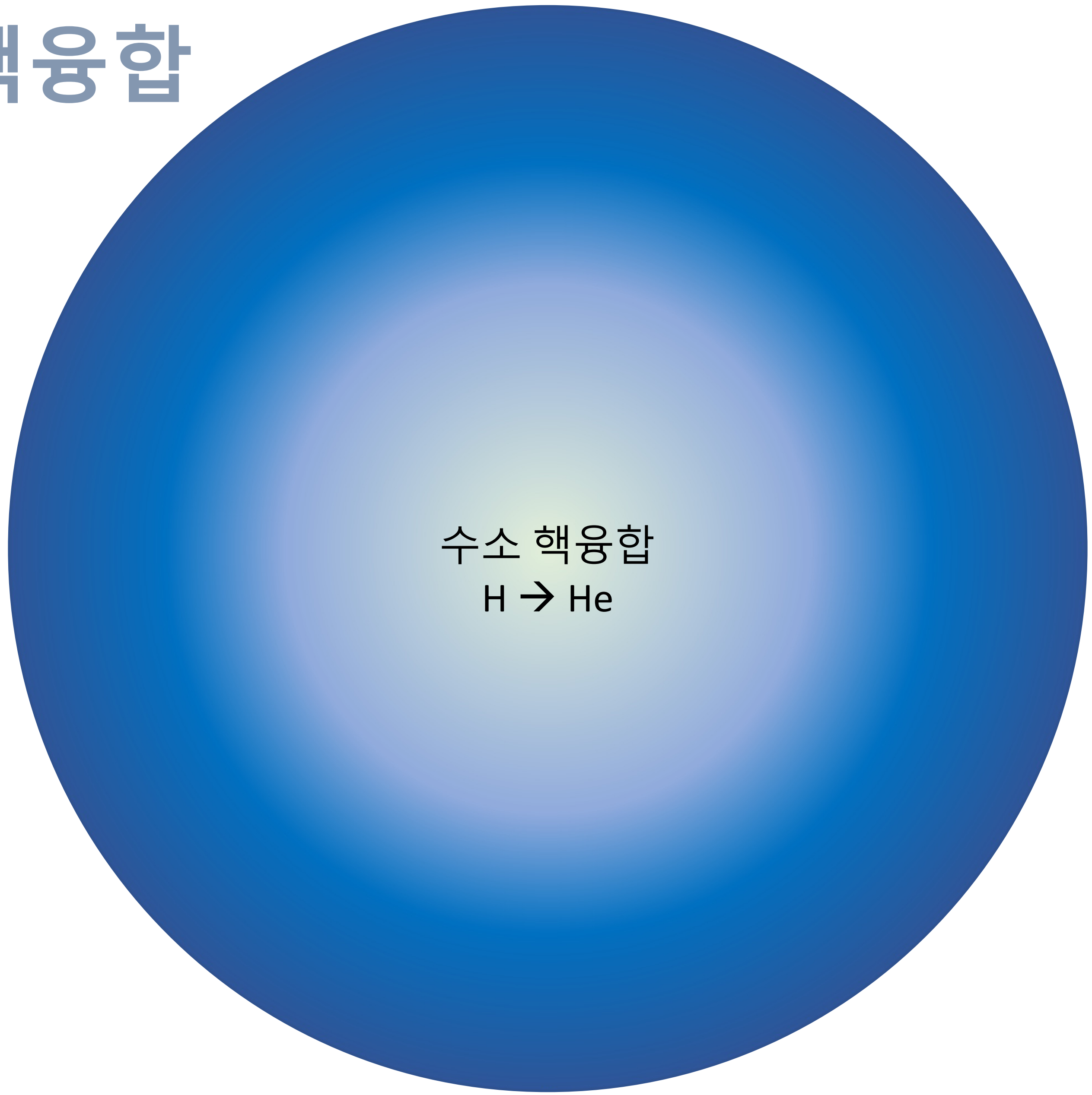


https://www.handprint.com/ASTRO/index.html

4. Overall Picture of Stellar Evolution

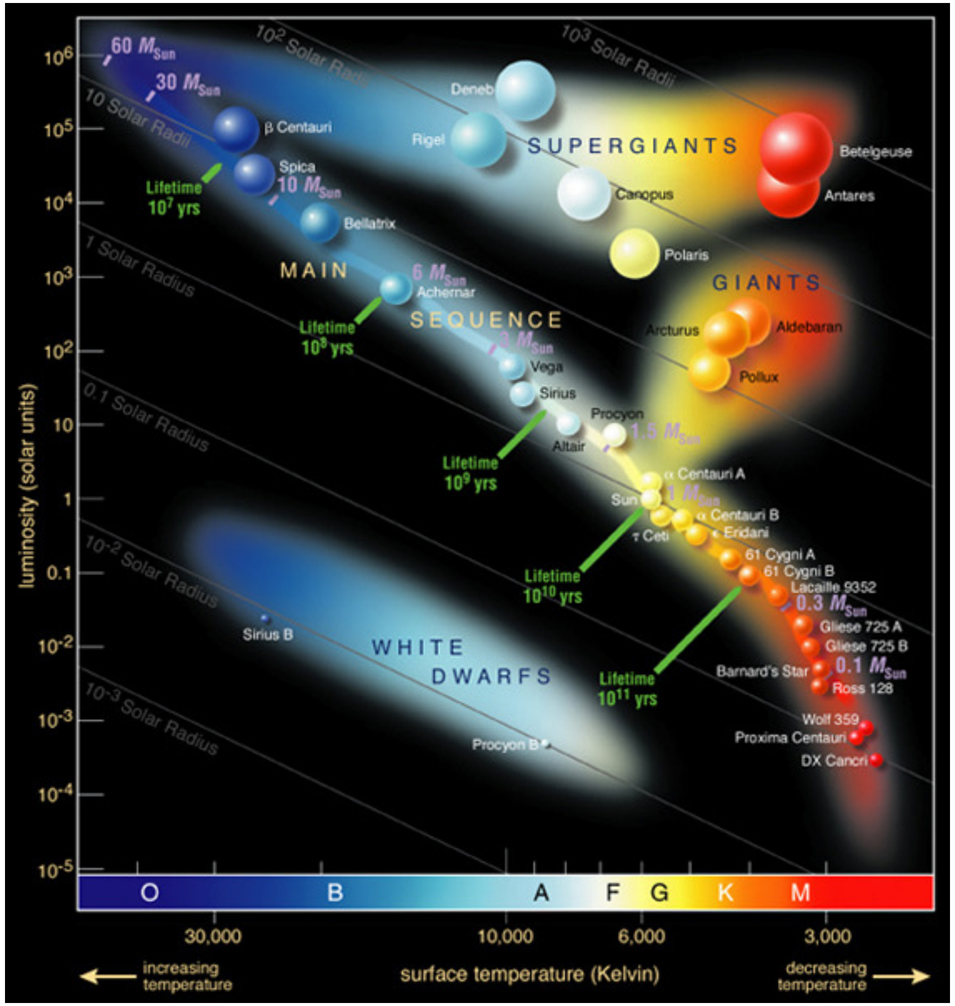
진화의 시작: 수소핵융합

주계열 : 수소핵융합 단계



Overall picture of stellar evolution

2. Main sequence

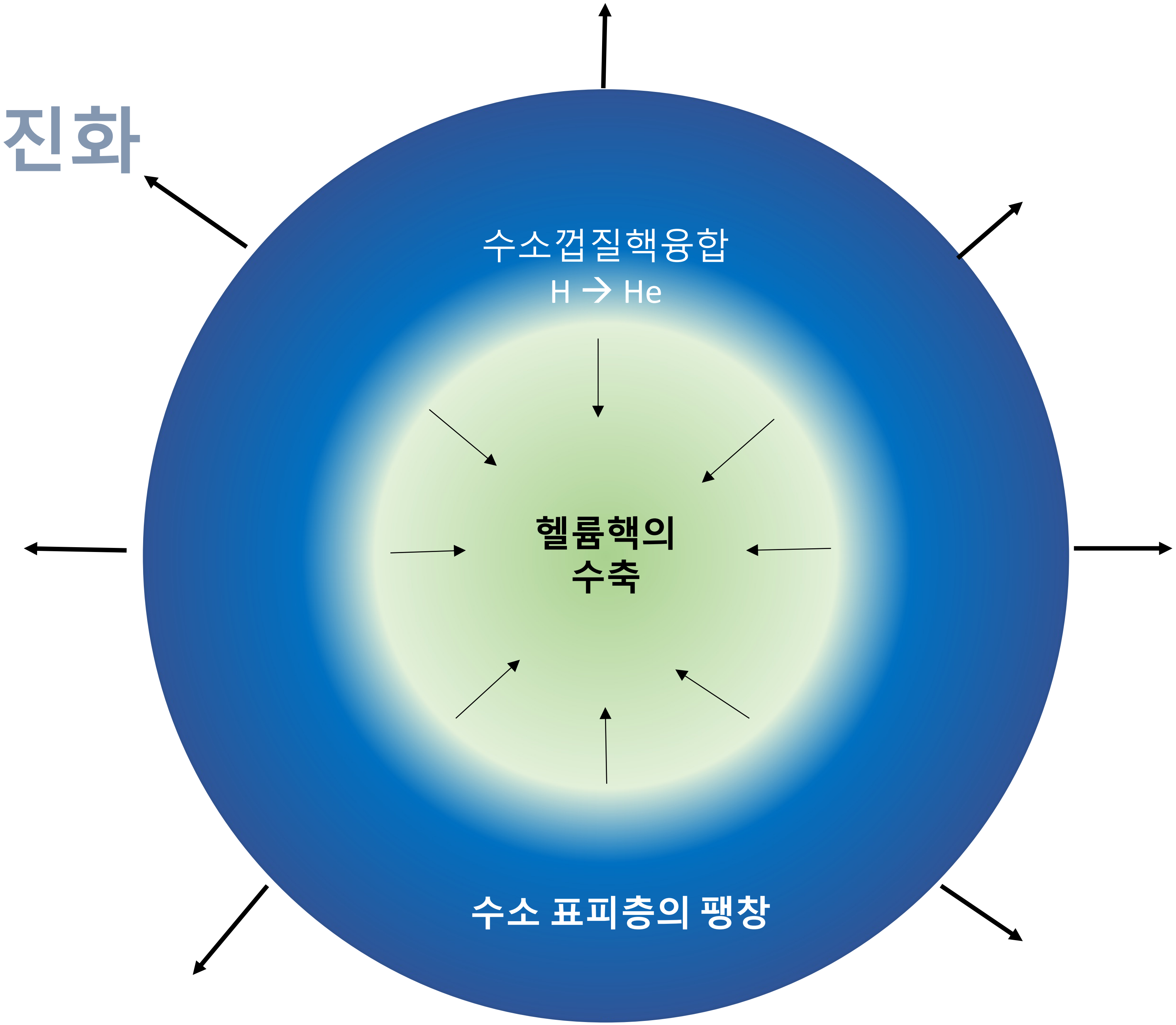


$$L \approx L_{\rm nuc}$$

on the main sequence

Stars on the main sequence
is in thermal equilibrium.

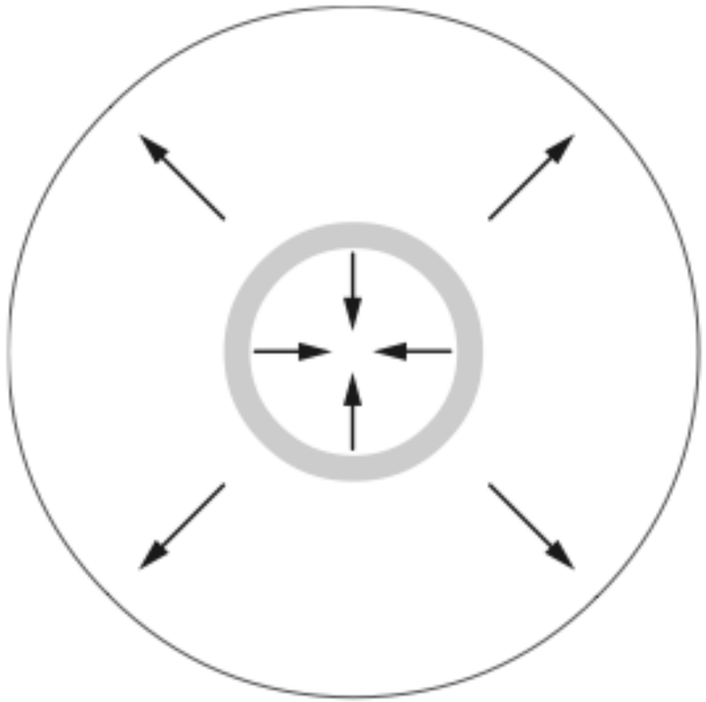
주계열 이후의 진화



Mirror principle

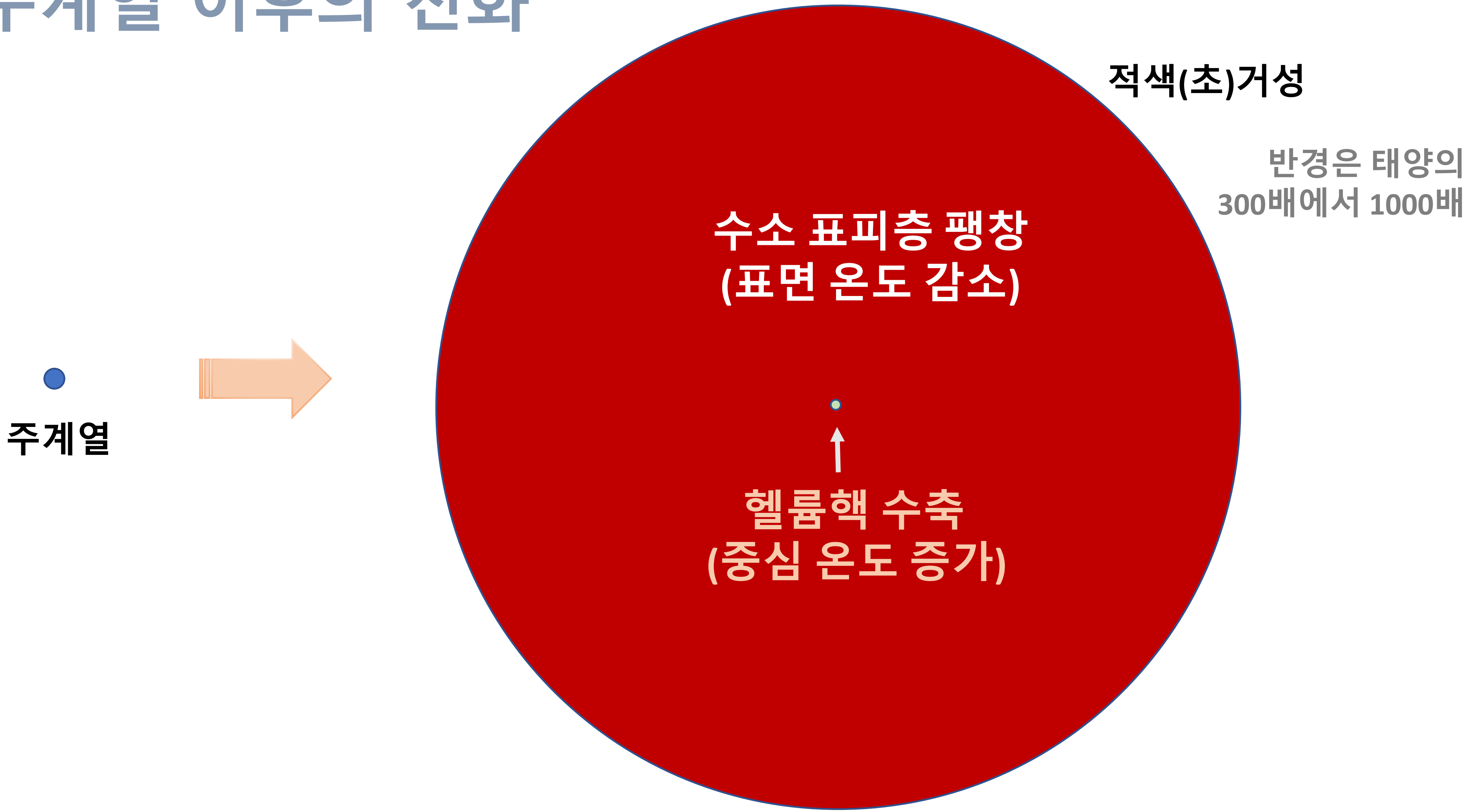
If a star has an active shell burning source, the burning shell acts as a mirror between the core and the envelope:

core contraction → envelope expansion
core expansion → envelope contraction



Numerical simulations show that, as the core contracts, the envelope expands, if there is a nuclear burning shell: I.e., the star becomes a giant star as the core contracts.

주계열 이후의 진화



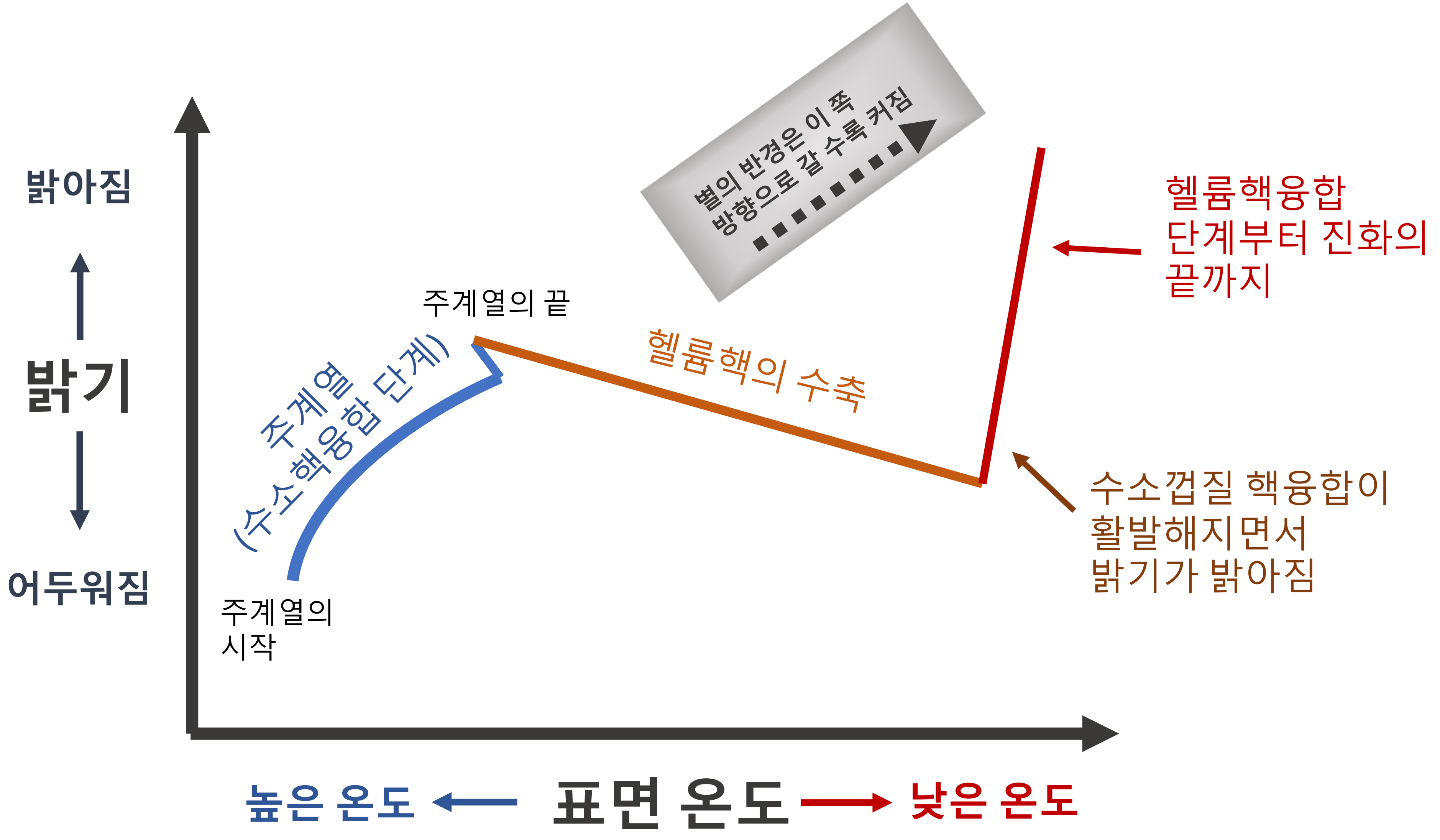
주계열 이후의 진화

헬륨핵융합 단계

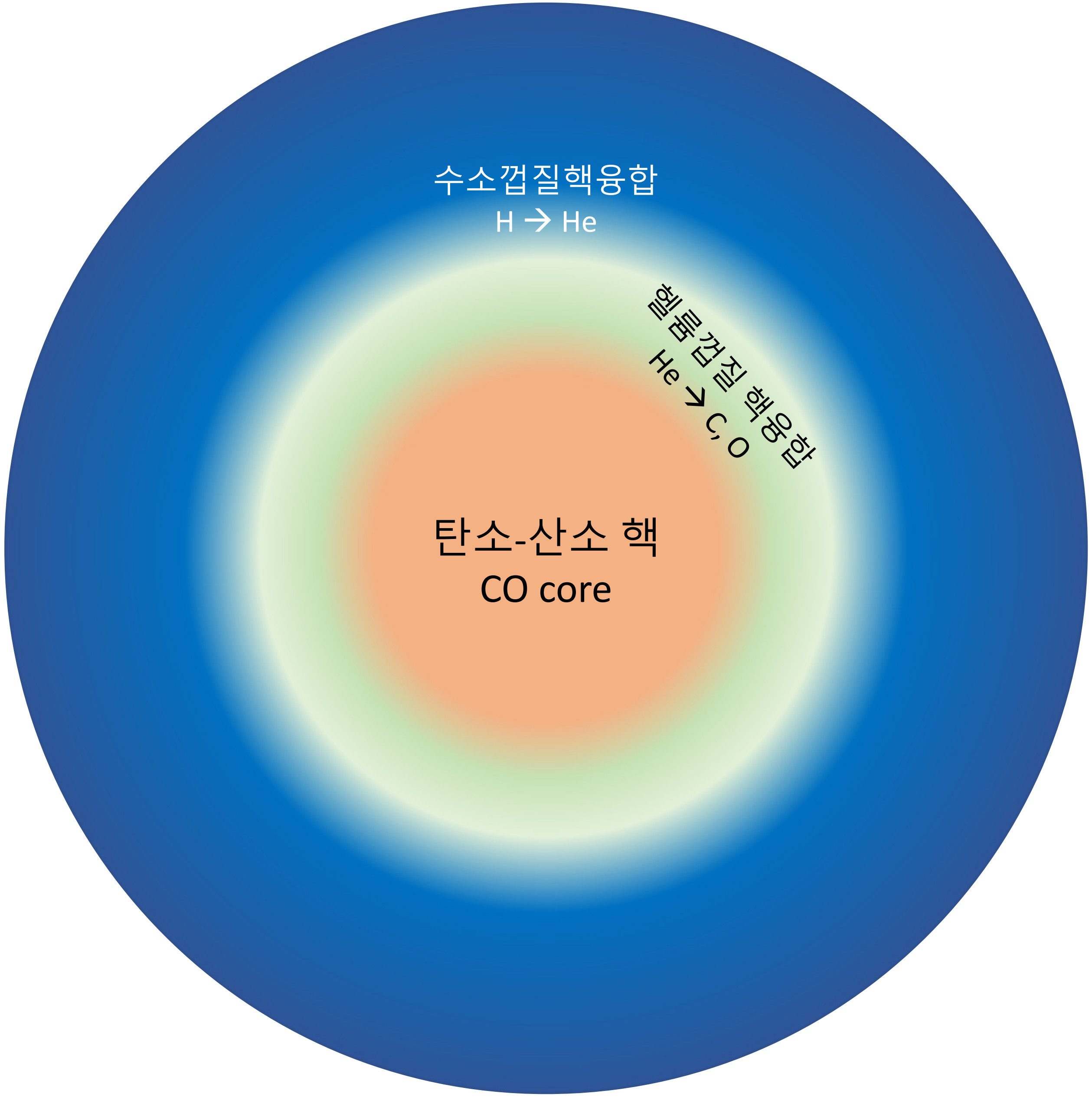
- **중심 온도: 1 - 2억 도**

$$3\,^{4}\mathrm{He} \longrightarrow {}^{12}\mathrm{C} + \gamma$$
$${}^{12}\mathrm{C} + {}^{4}\mathrm{He} \longrightarrow {}^{16}\mathrm{O} + \gamma$$

진화에 따른 밝기와 표면 온도 변화 (M > ~ 2 Msun 인 경우)



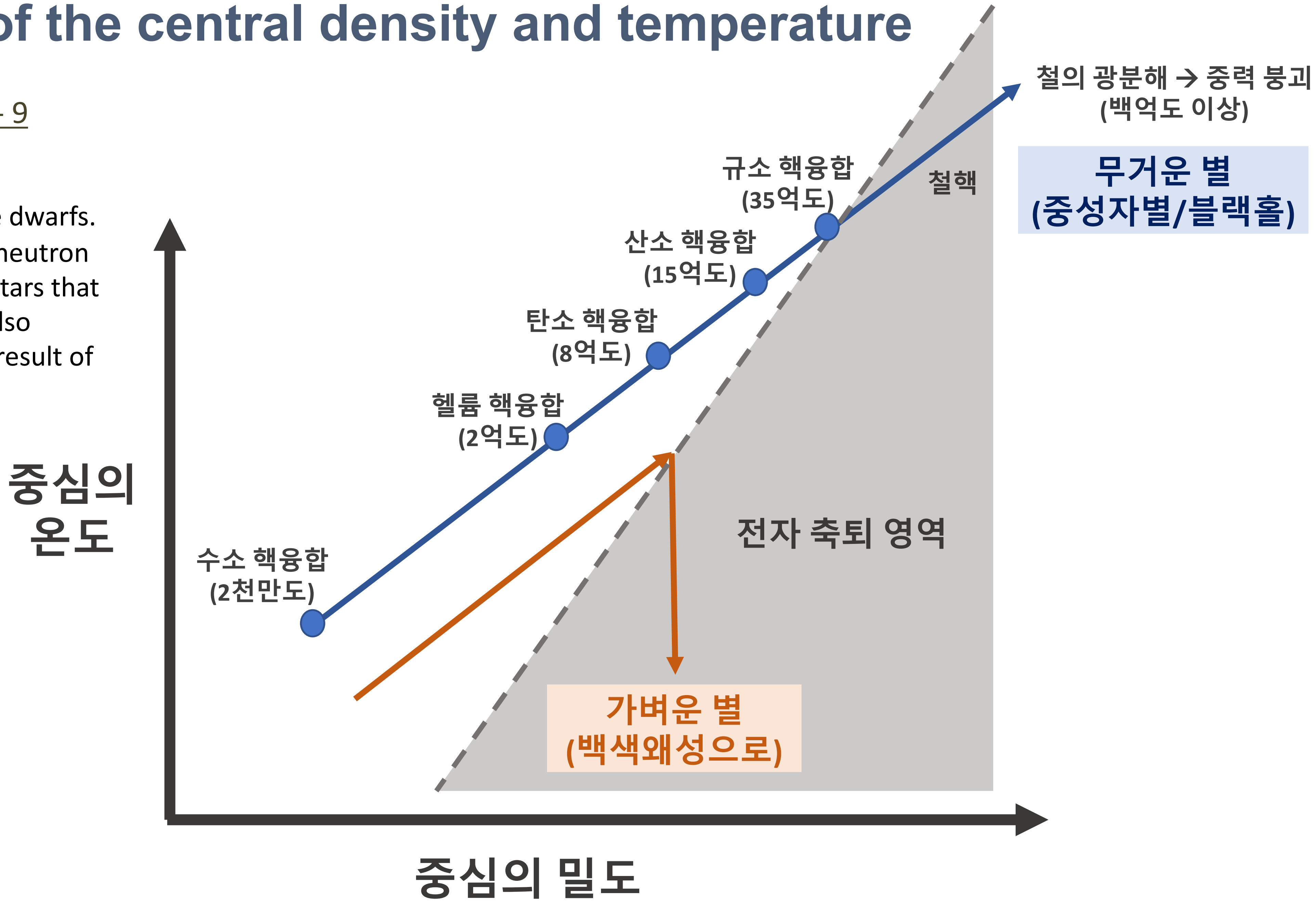
주계열 이후의 진화



Evolution of the central density and temperature

Stars with M > about 8 - 9 Msun

They do not become white dwarfs. The iron cores collapse to neutron stars or black holes. The stars that collapse to neutron stars also produce supernovae, as a result of the core-collapse.



II. Evolution of Massive Stars

https://commons.wikimedia.org/wiki/File:Night_Sky_looking_towards_Orion.jpg

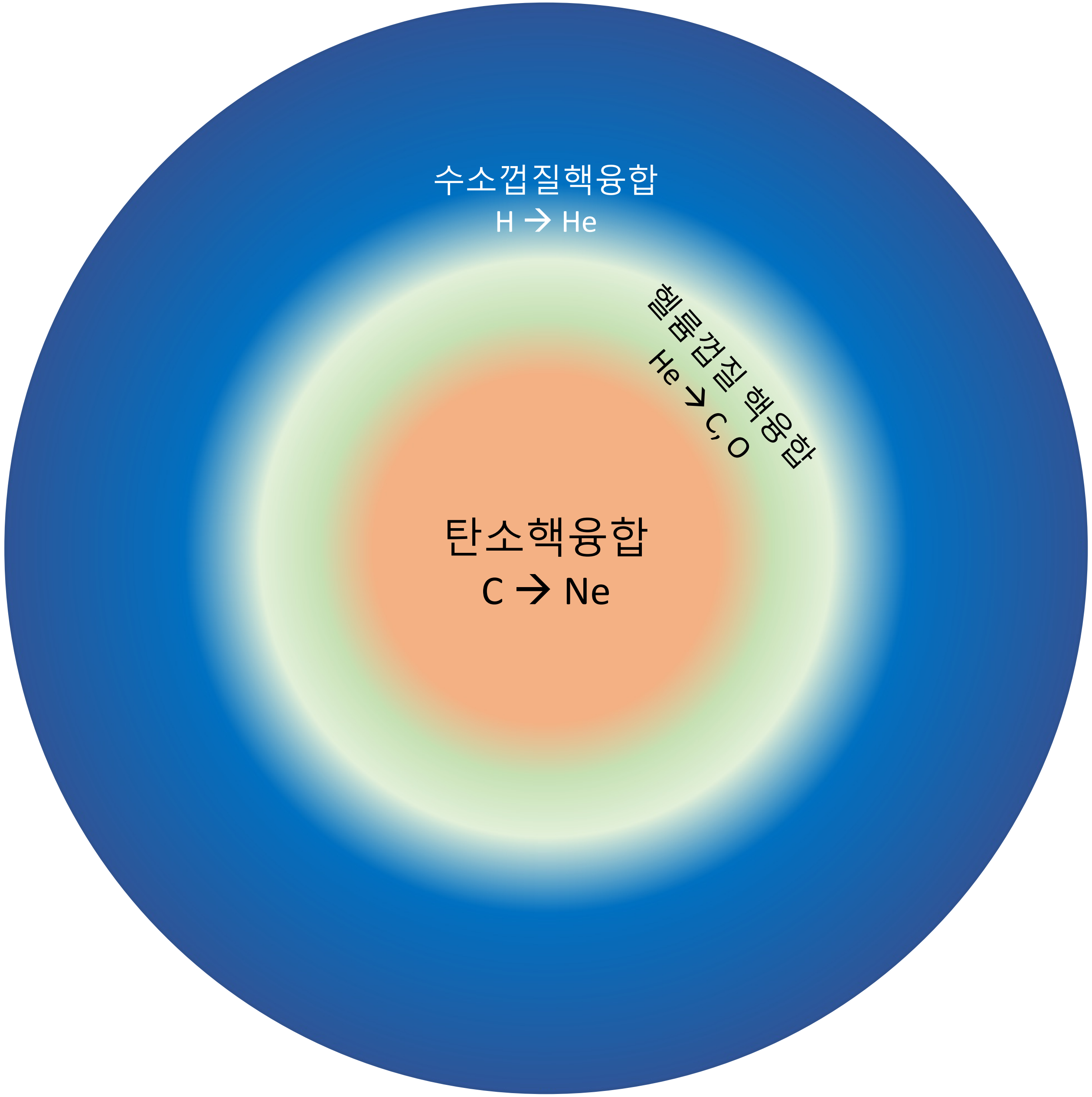
주계열 이후의 진화

태양보다 20배 질량이 큰 경우의 예:

탄소핵융합 단계

- 중심 온도: 8억 도
- 지속시간: 천 년

$$2\,^{12}\mathrm{C} \longrightarrow {}^{20}\mathrm{Ne} + {}^{4}\mathrm{He}$$
$$\longrightarrow {}^{24}\mathrm{Mg} + \gamma$$



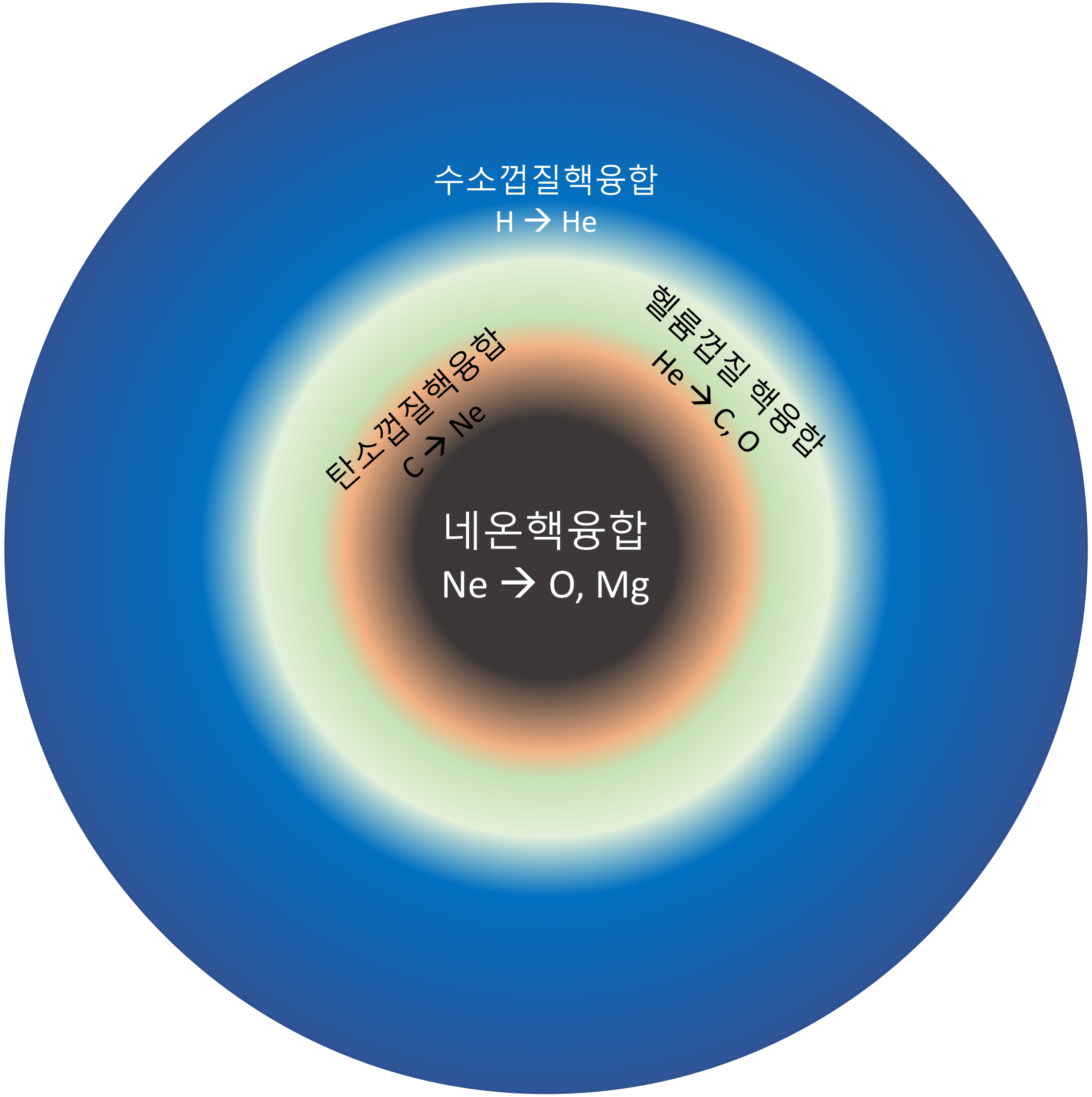
주계열 이후의 진화

태양보다 20배 질량이 큰 경우의 예:

네온핵융합 단계

- 중심 온도: 15억 도
- 지속시간: 3 년

$$^{20}\mathrm{Ne} + \gamma \longrightarrow {}^{16}\mathrm{O} + {}^{4}\mathrm{He}$$
$$^{20}\mathrm{Ne} + {}^{4}\mathrm{He} \longrightarrow {}^{24}\mathrm{Mg} + \gamma$$



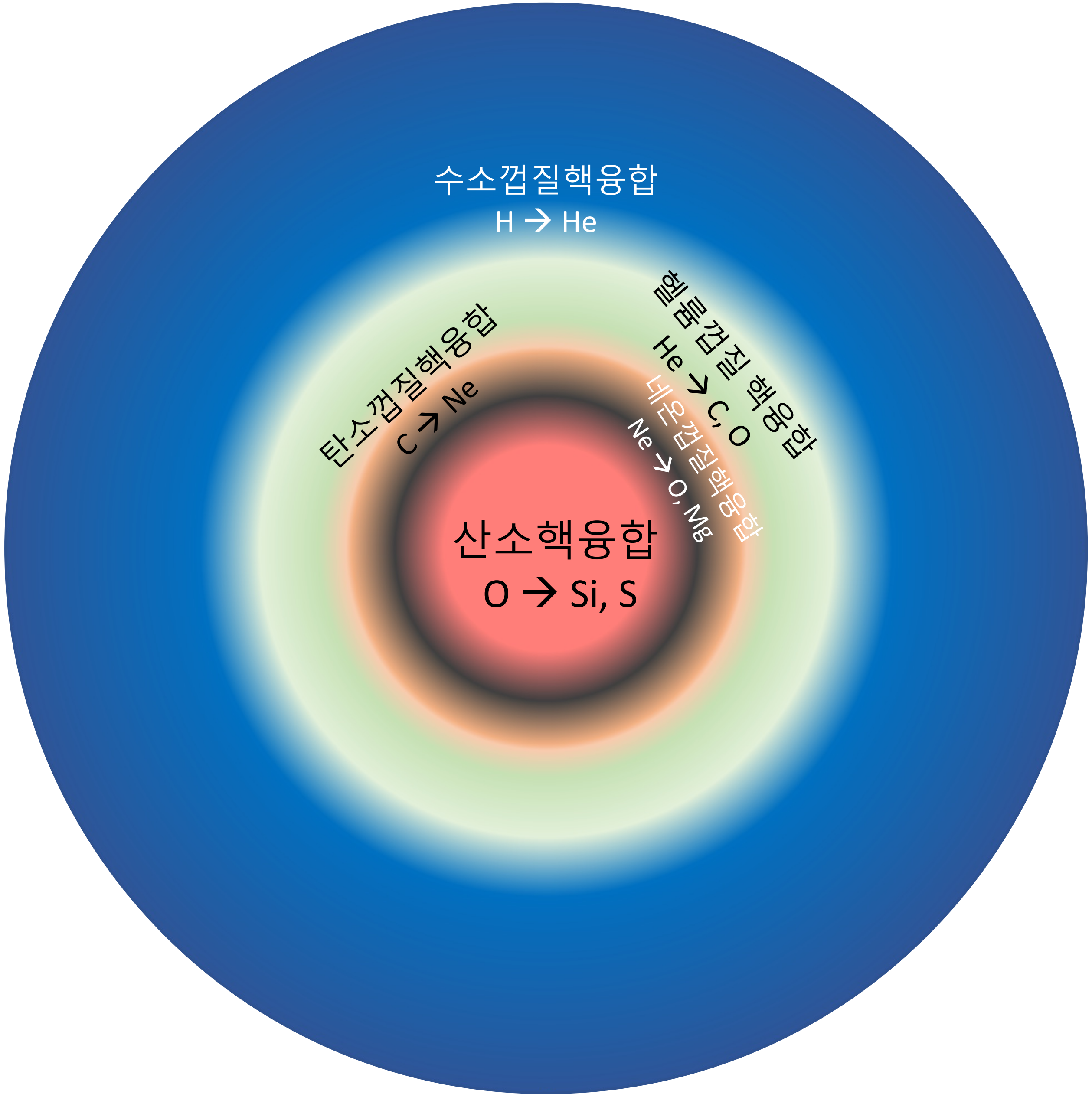
주계열 이후의 진화

<u>**태양보다 20배 질량이 큰 경우의 예:**</u>

산소핵융합 단계

- 중심 온도: 20억 도
- 지속시간: 10개월

$$2\,^{16}\mathrm{O} \longrightarrow {}^{28}\mathrm{Si} + {}^{4}\mathrm{He}$$
$$\longrightarrow {}^{32}\mathrm{S} + \gamma$$

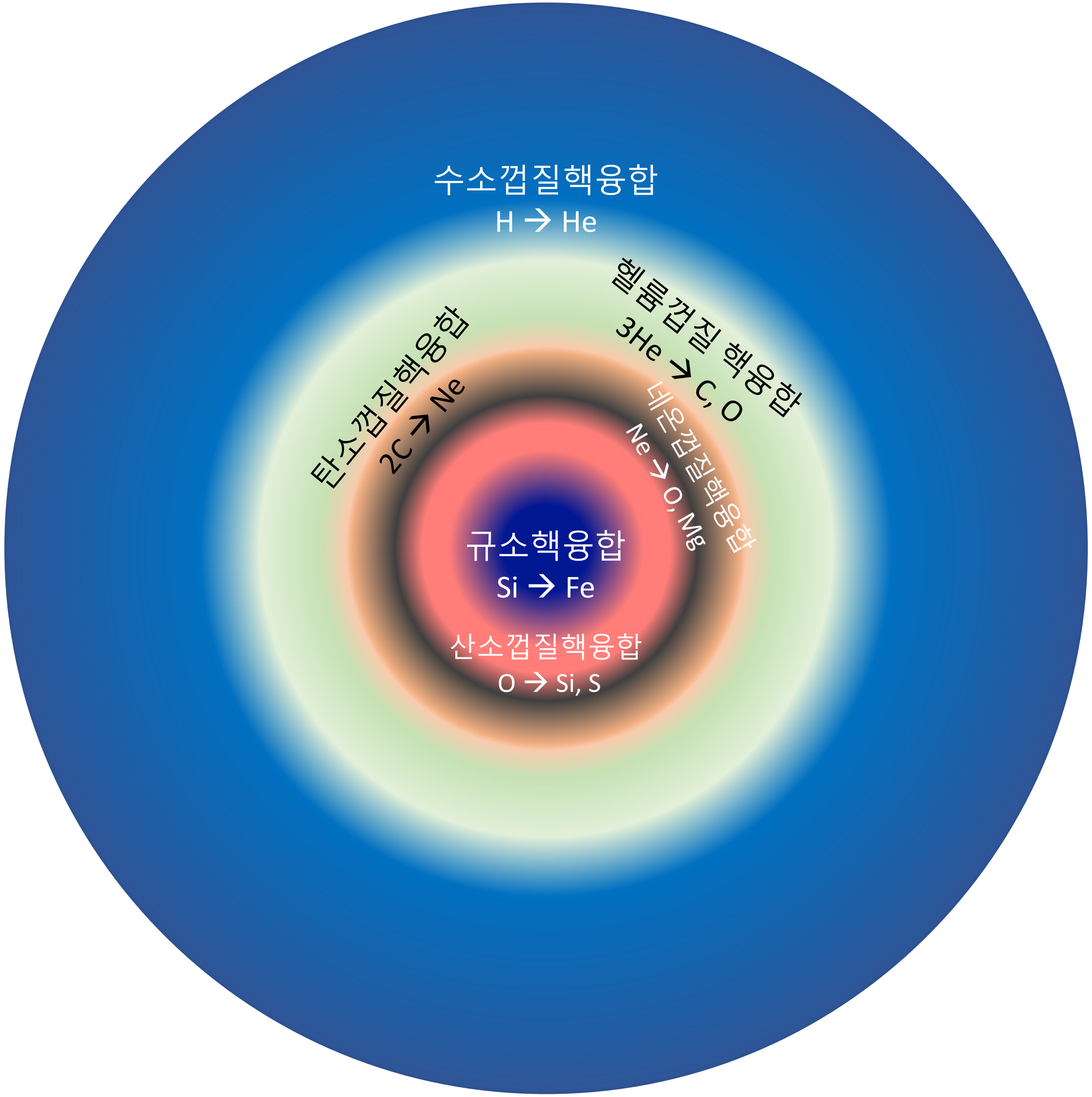


주계열 이후의 진화

태양보다 20배 질량이 큰 경우의 예:

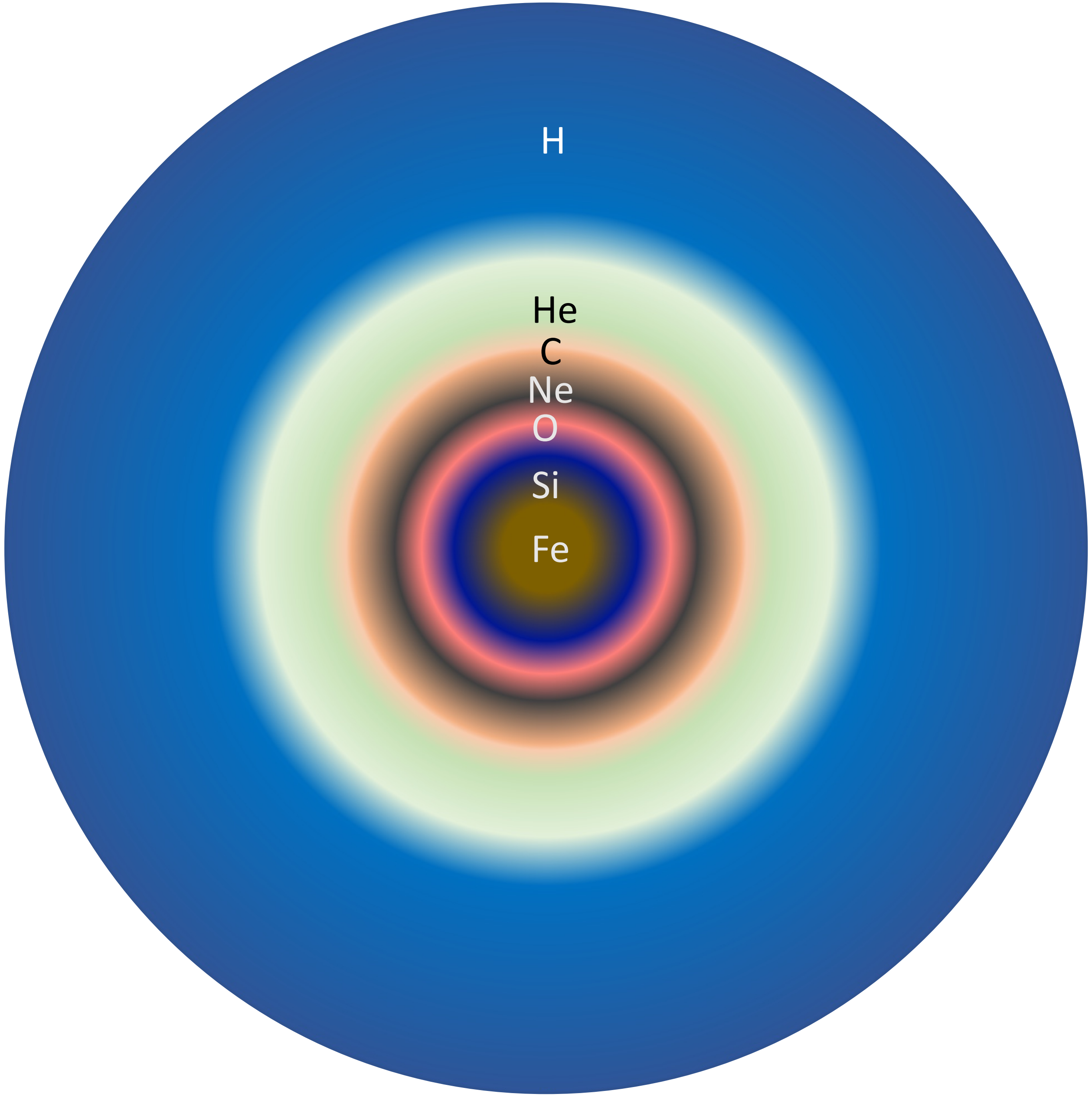
규소핵융합 단계

- 중심 온도: 35억 도
- 지속시간: 1주일

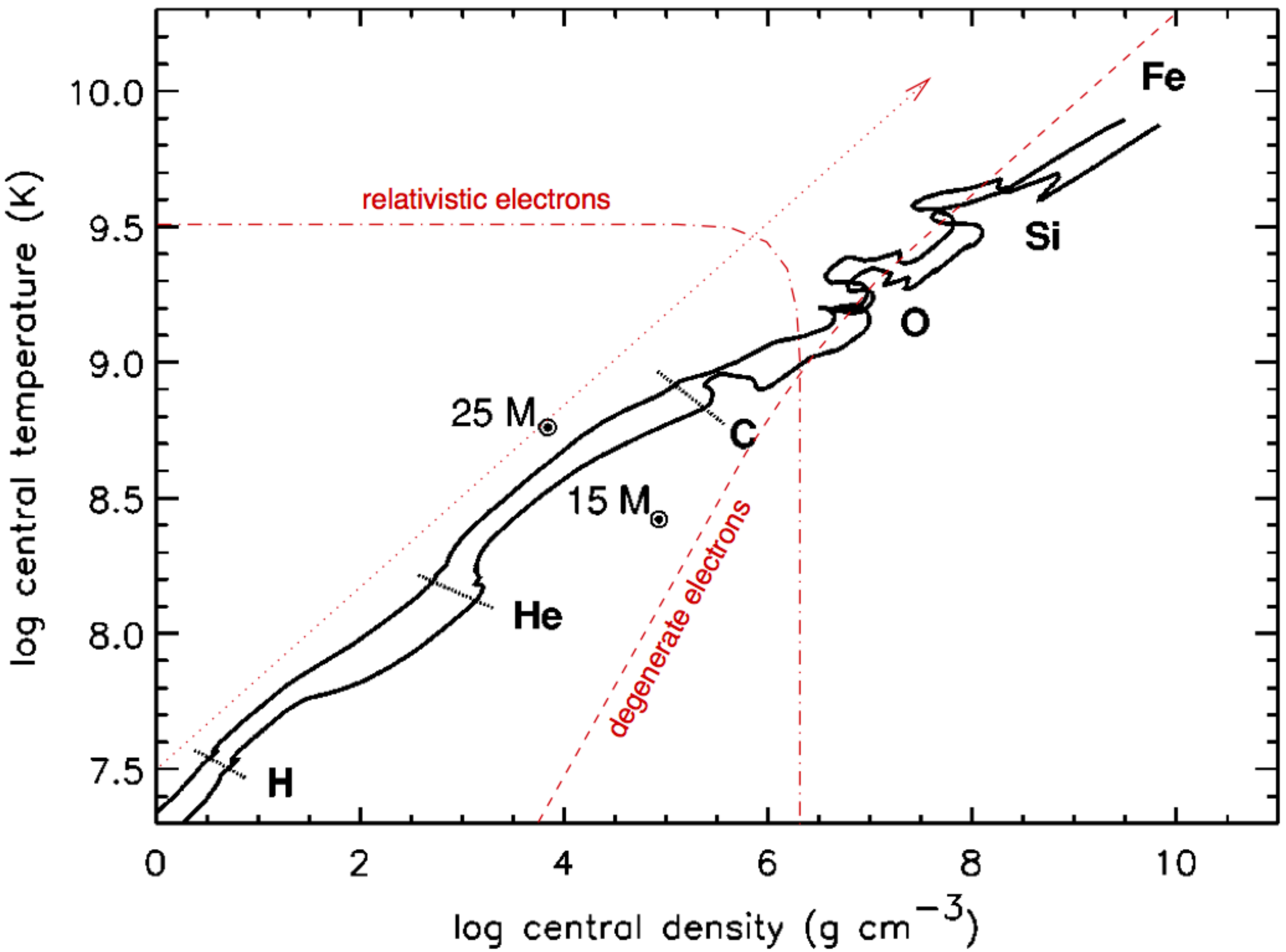


주계열 이후의 진화

진화의 최종 단계 모습: 양파 구조



Evolution of Massive Stars



Advanced Nuclear Burning Stages

(e.g., 20 solar masses)

Fuel	Main Product	Secondary Products	Temp (10^9 K)	Time (yr)
H	He	^{14}N	0.02	10^7
He	C,O	^{18}O, ^{22}Ne s- process	0.2	10^6
C	Ne, Mg	Na	0.8	10^3
Ne	O, Mg	Al, P	1.5	3
O	Si, S	Cl, Ar K, Ca	2.0	0.8
Si	Fe	Ti, V, Cr Mn, Co, Ni	3.5	1 week

Courtesy: S. Woosly

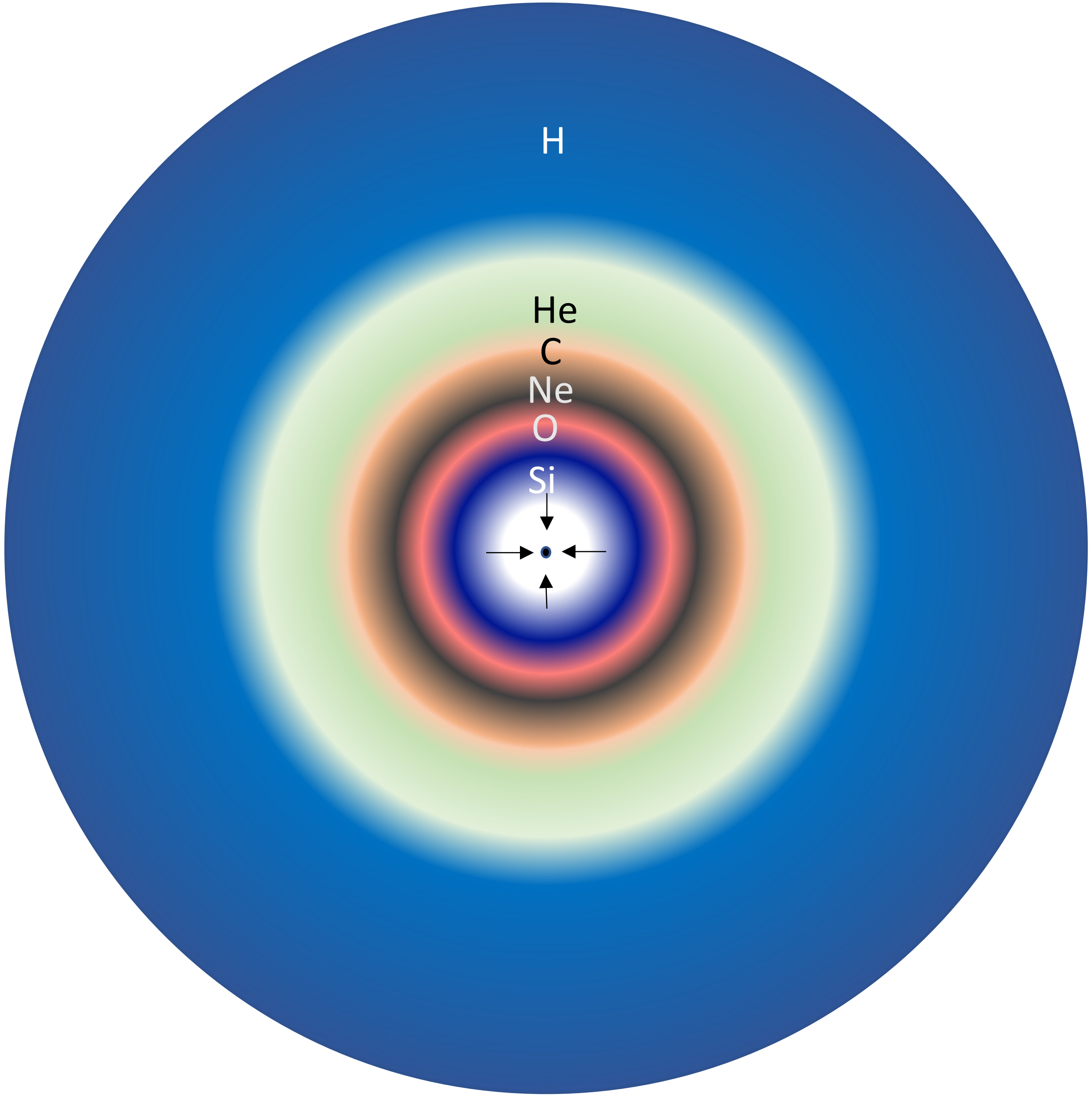
무거운 별의 죽음

철의 광분해 (photodissociation)와 그에 따른 철핵(iron core)의 중력 붕괴

$$Fe \rightarrow p,\ n,\ e^-$$

$$p + e^- \rightarrow n + \nu$$

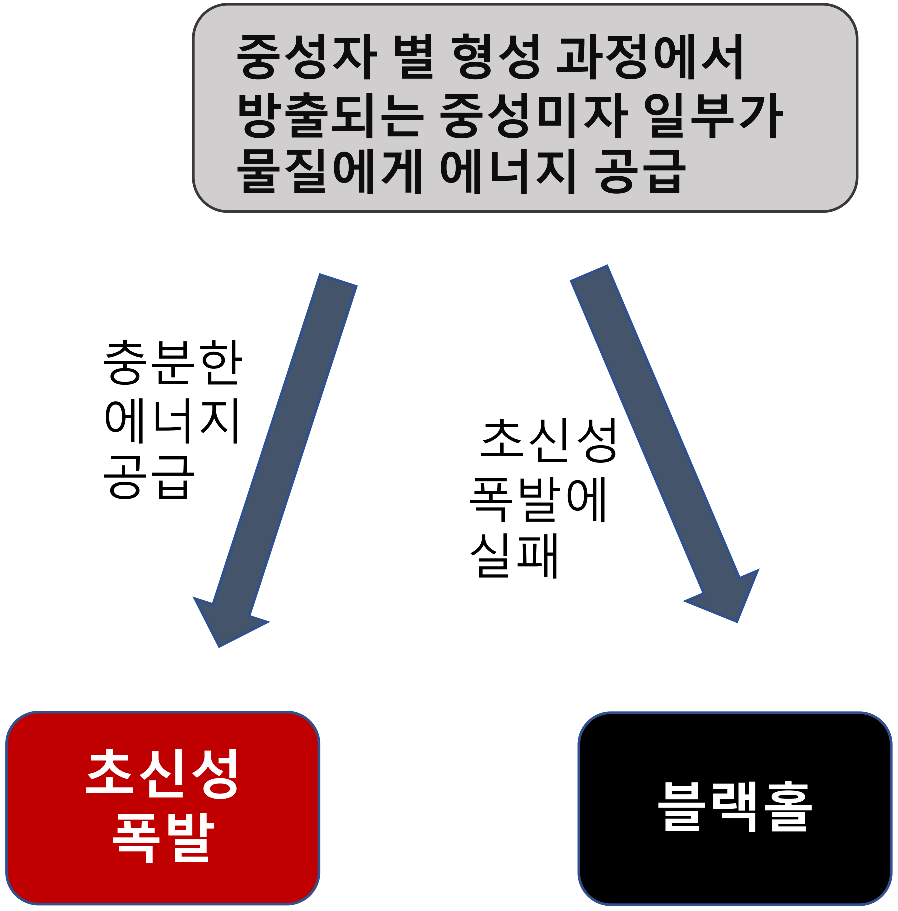
이렇게 생성된 중성자들로 구성이 된 중성자 별이 탄생.

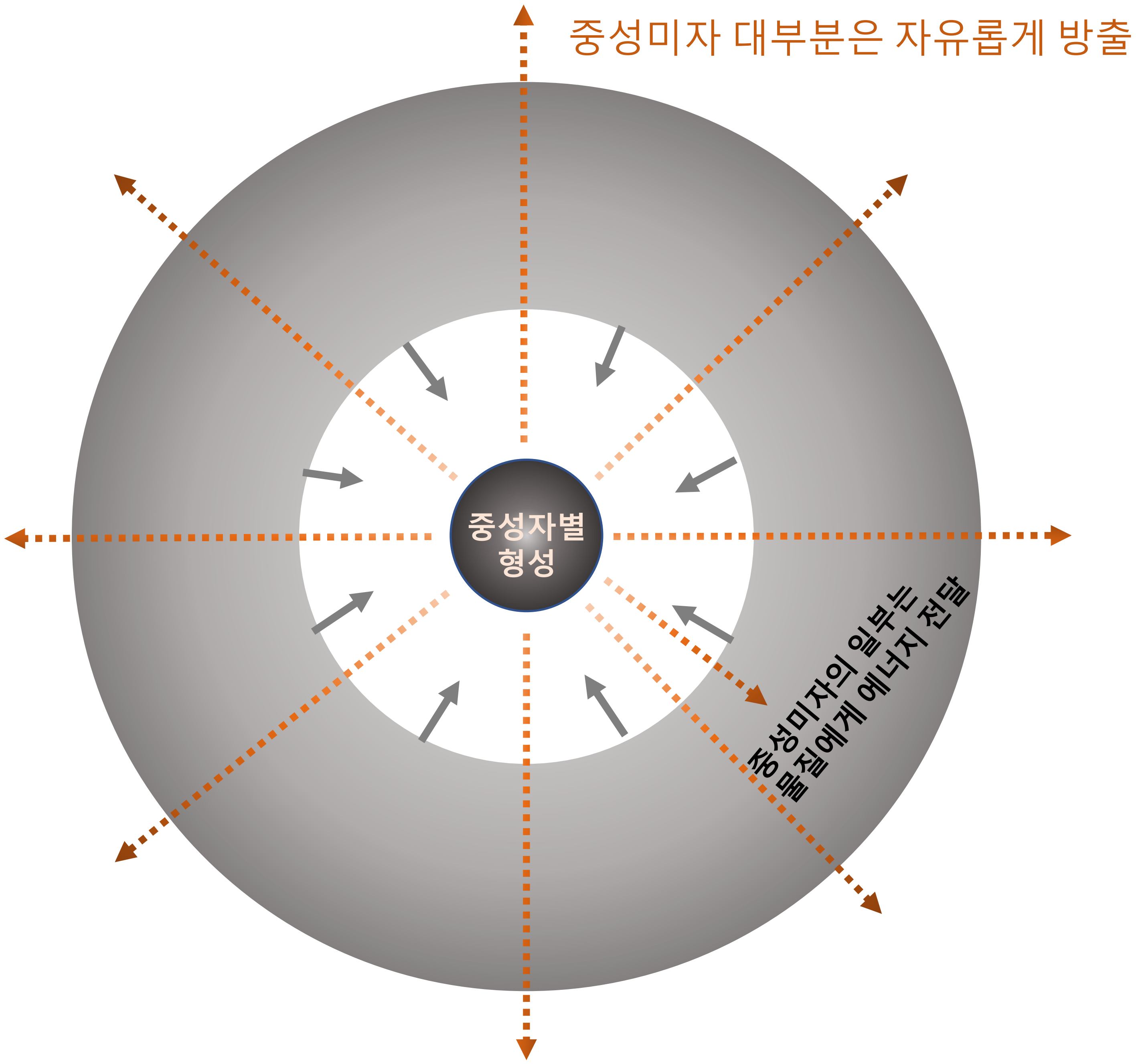


중성자별

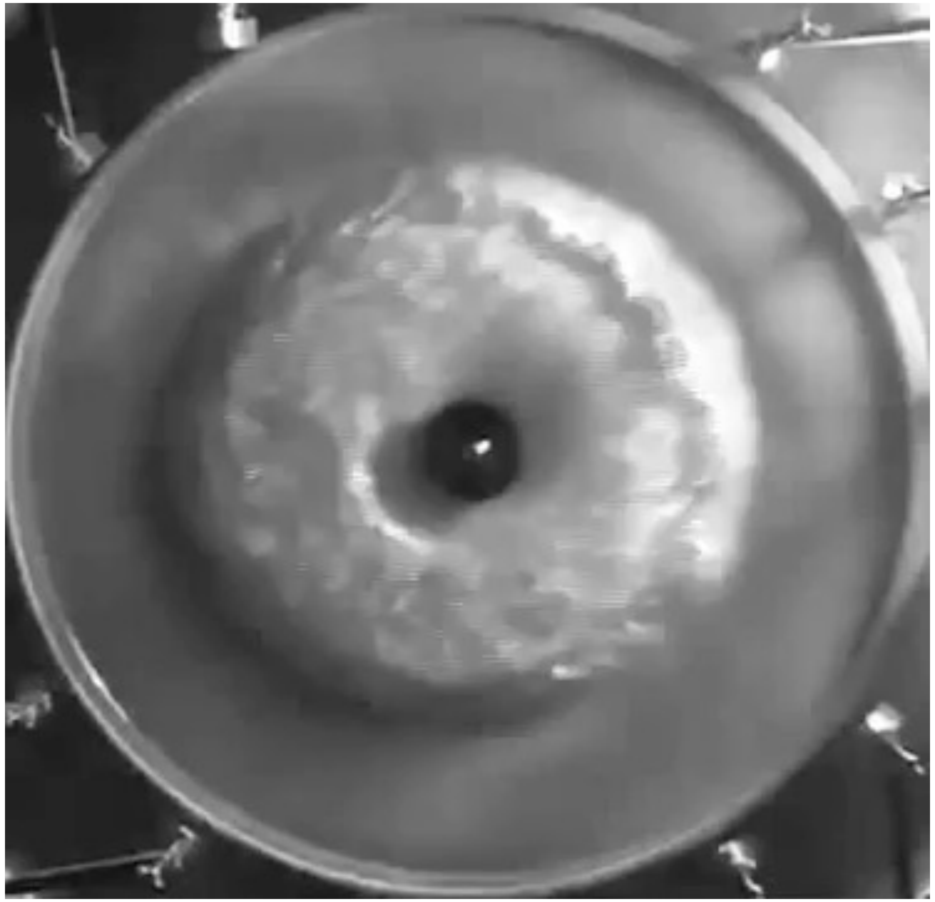
반경 10km의 공간에 지구 질량의 50만 배, 혹은 태양 질량의 1.4 배가 응축되어 있는 고밀도 천체

초신성 폭발 메커니즘


중성자 별 형성 과정에서
방출되는 중성미자 일부가
물질에게 에너지 공급
충분한
에너지
공급
초신성
폭발에
실패
초신성
폭발
블랙홀


중성미자 대부분은 자유롭게 방출
중성자별
형성
중성미자의 일부는
물질에게 에너지 전달

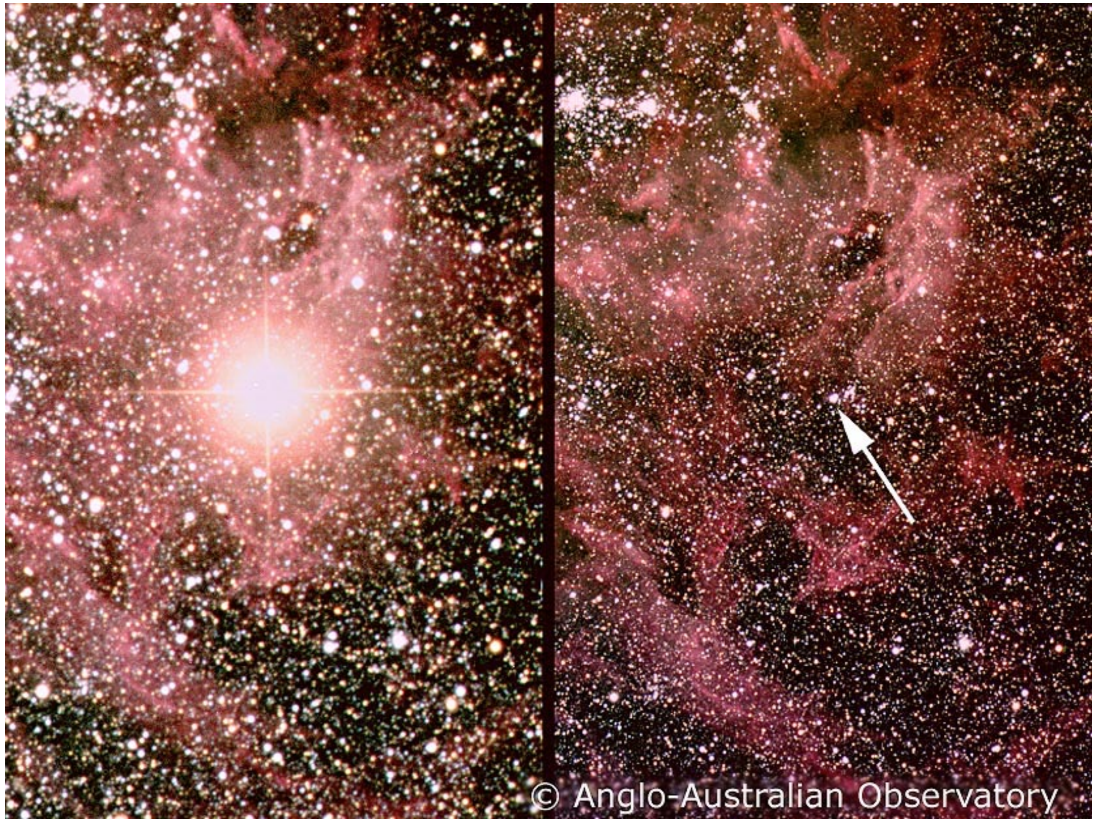
참고: SASI: Standing Accretion Shock Instability

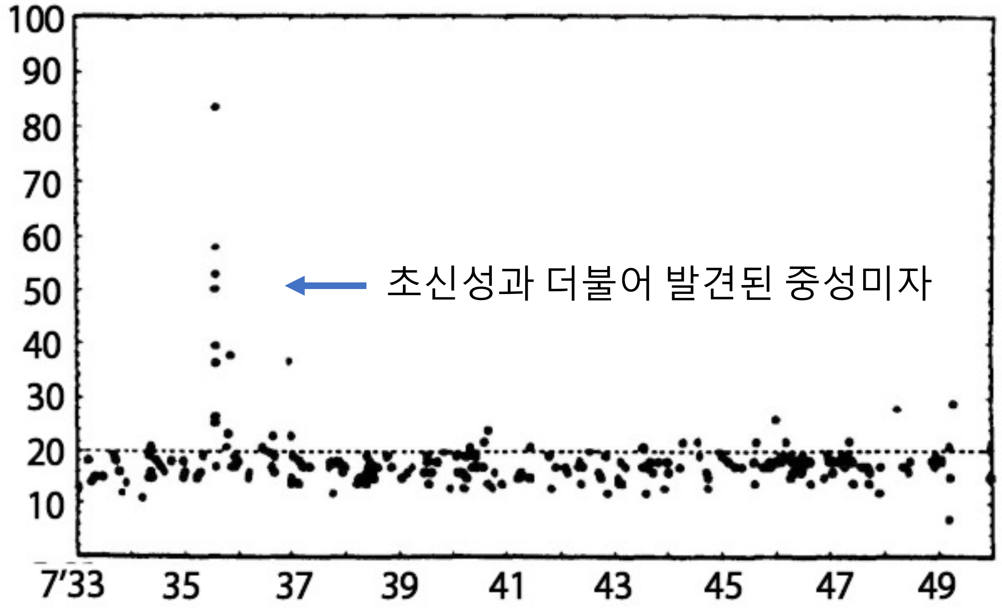


https://www.youtube.com/watch?v=5fcsSA31rkE

https://physics.aps.org/articles/v5/13

초신성 1987A: 무거운 별이 초신성을 만든다는 최초의 직접적인 증거

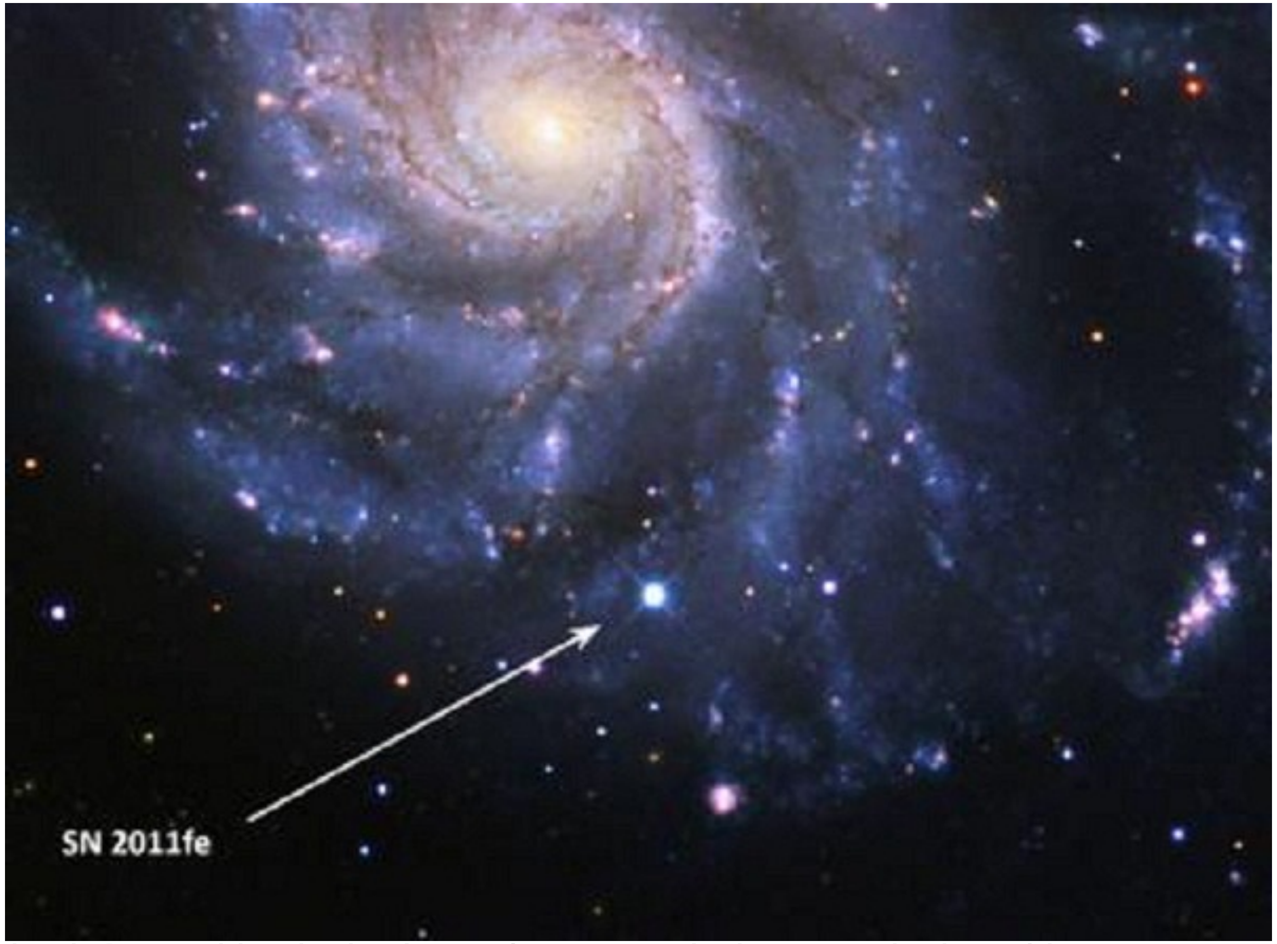




https://www.u-tokyo.ac.jp/focus/en/features/f_00038.html

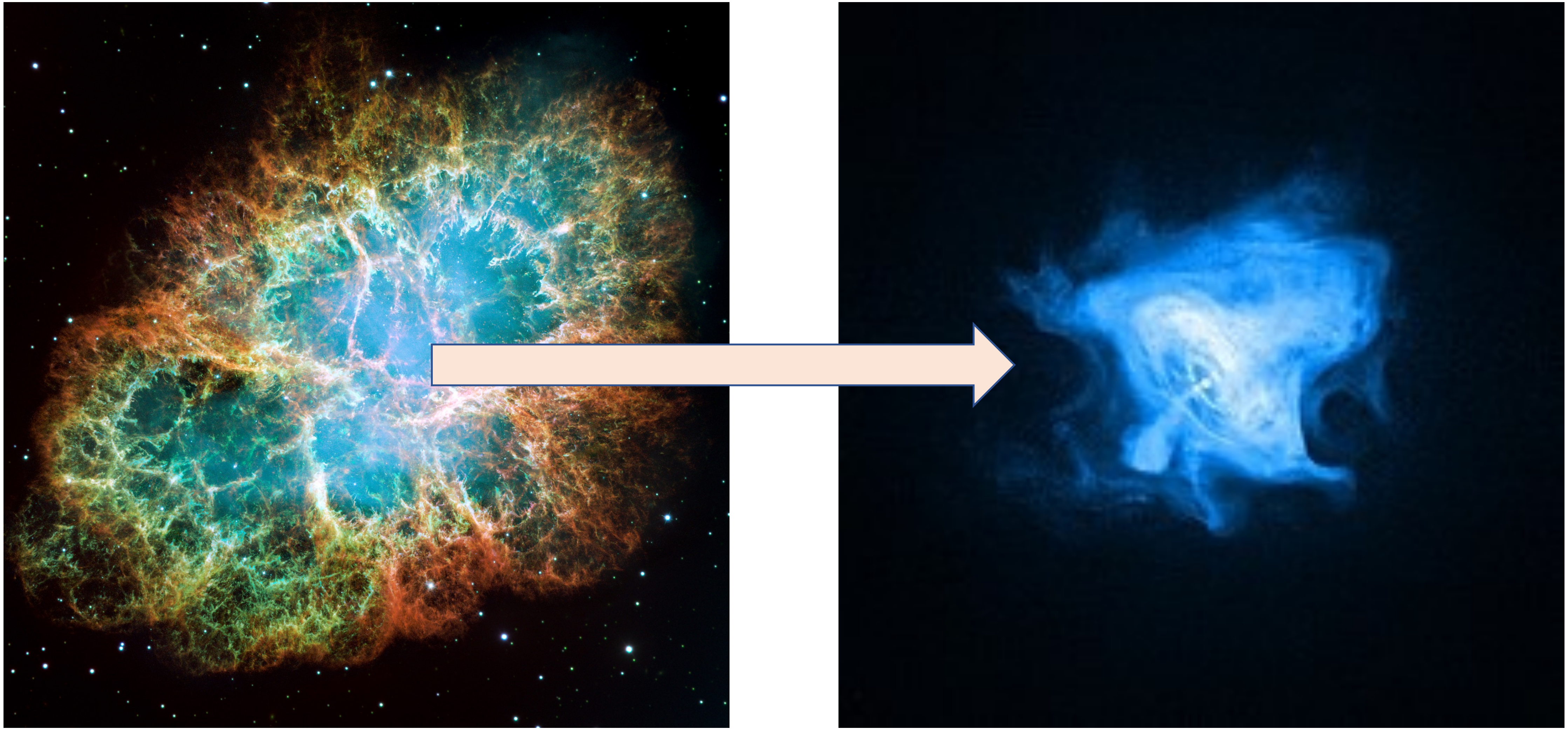
왼쪽: 초신성 폭발 장면, 오른쪽: 초신성 폭발 이전에 찍힌 초신성 모체성의 사진

초신성폭발

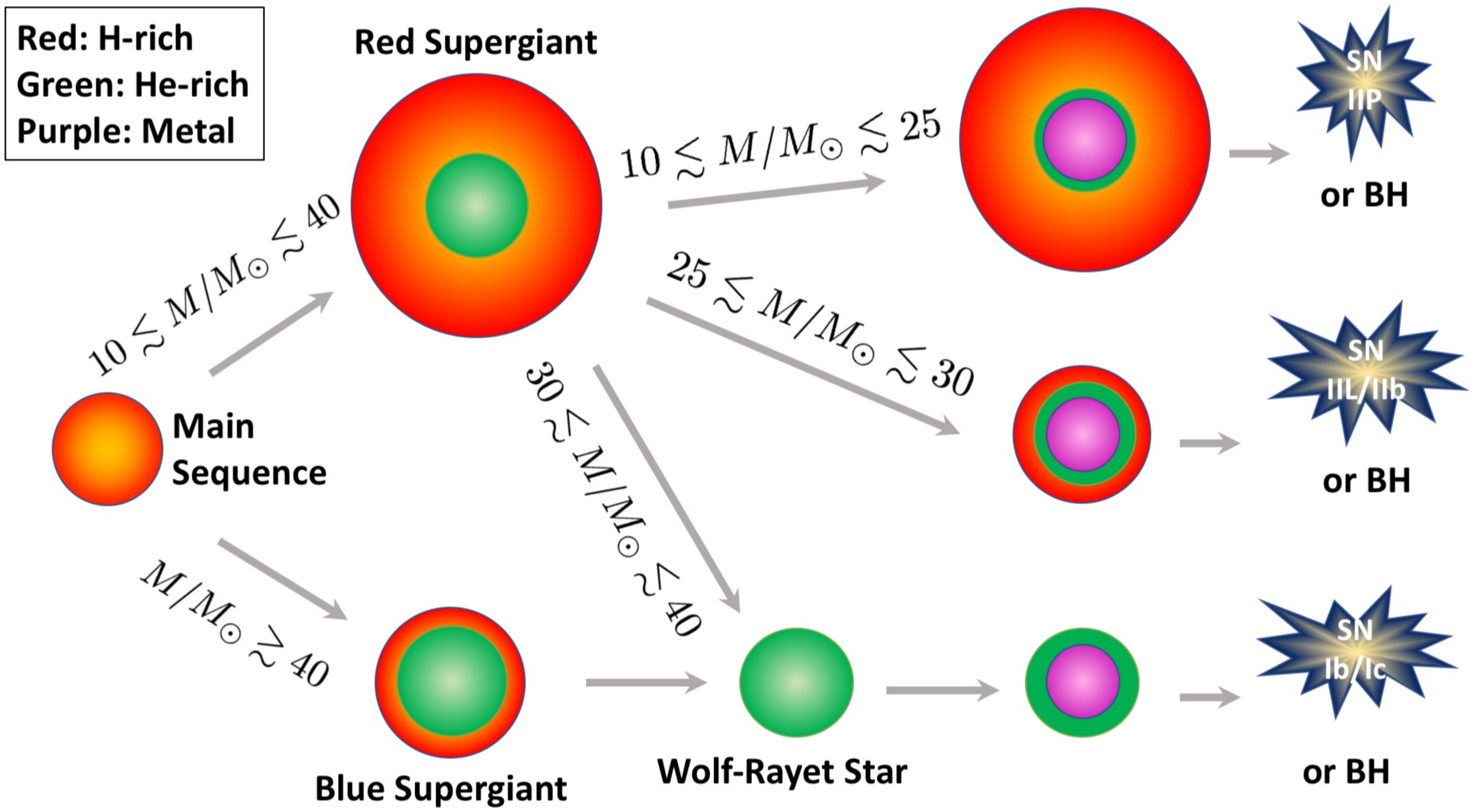


초신성은 태양보다 억 배에서 수천억 배가 밝음. 은하 전체 만큼, 혹은 그 이상으로 밝기도 함.

초신성 잔해에서 발견되는 중성자별 / 펄사



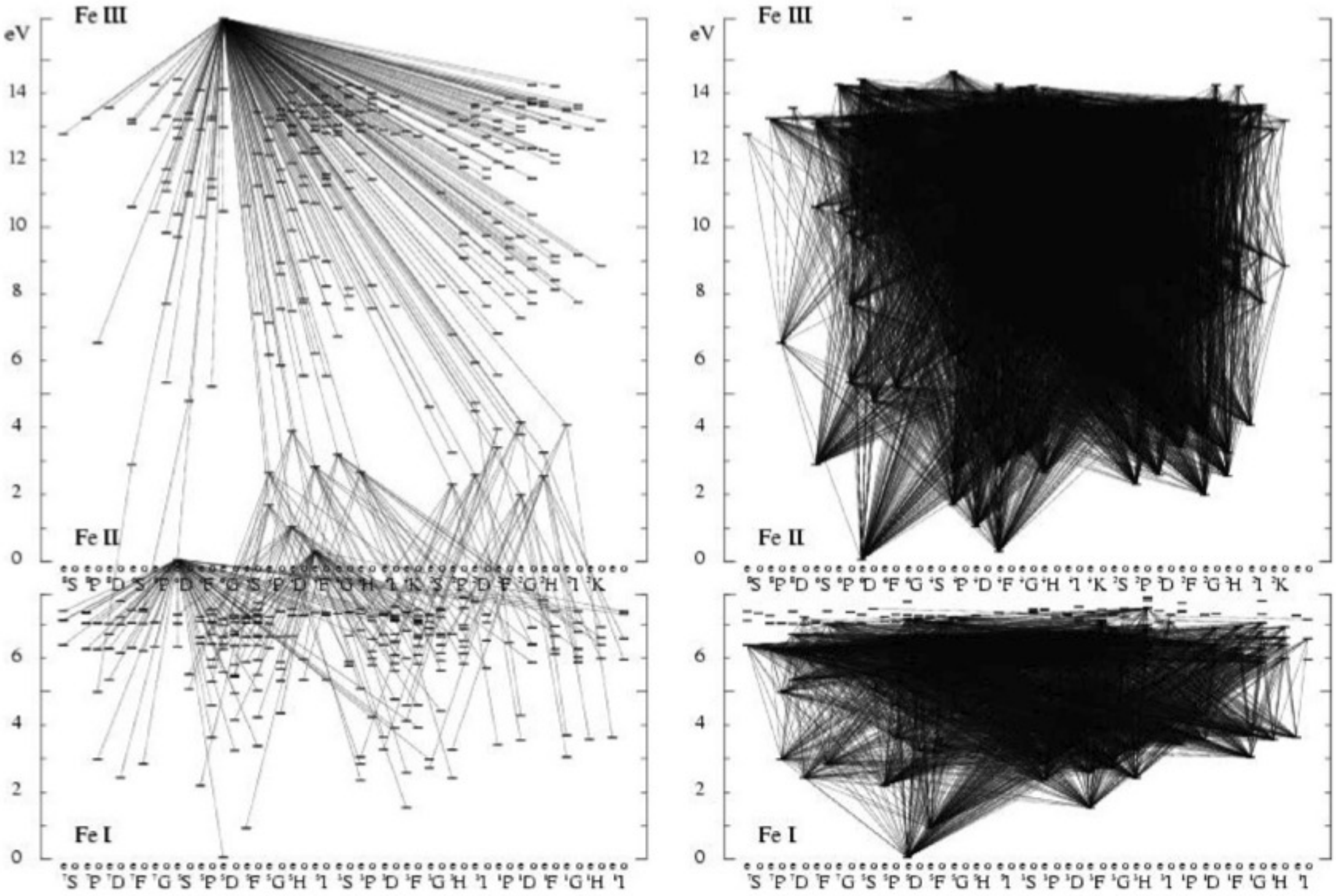
Standard scenario of massive star evolution



$$g_{\rm rad} = \int \frac{\kappa_\nu F_\nu}{c} d\nu$$

$$\dot{M}_{\rm wind} \propto L^{1.5} Z^{0.7}$$

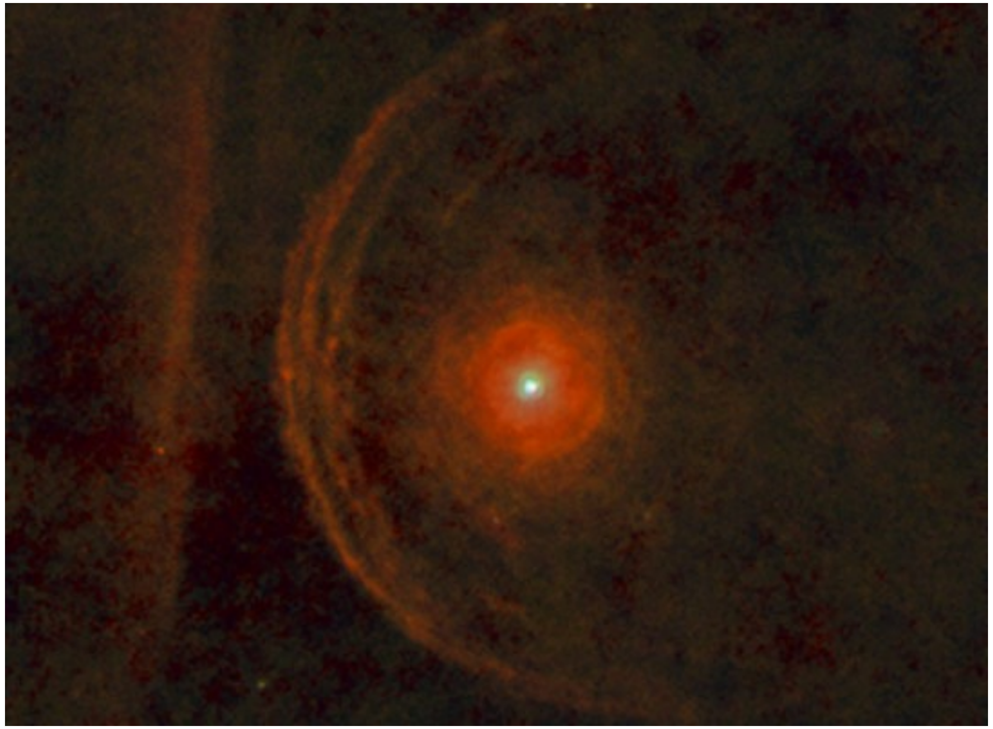
Iron gives almost a grey opacity.



Gehren et al. 2001; *left*: bf transtion, *right*: bb transtion

Red-supergiant star: Betelgeuse

This is the closest RSG from the earth (about 640 light year away)

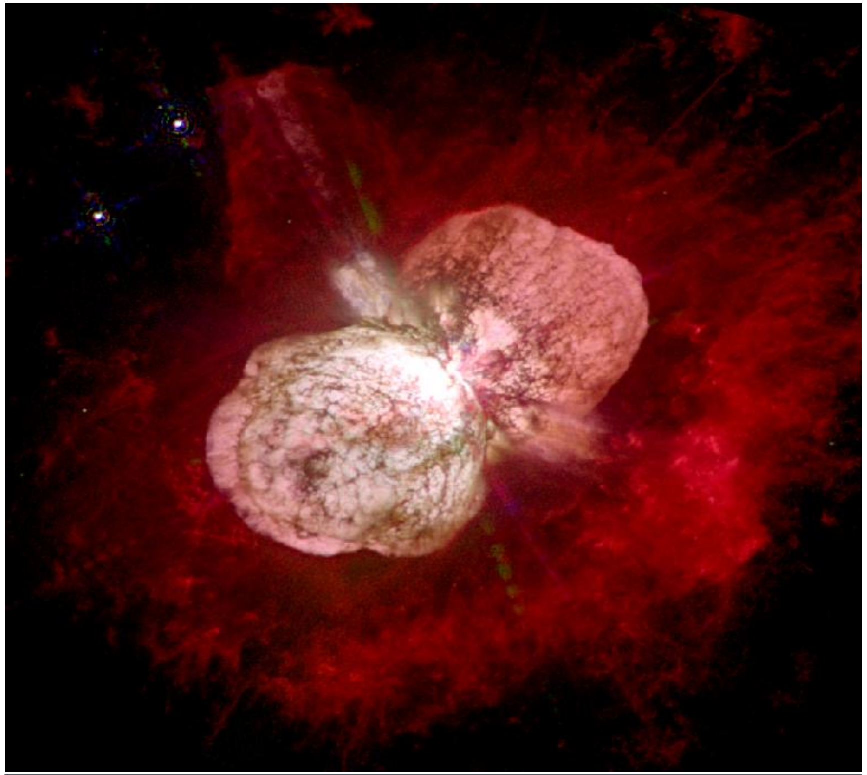


IR image by the Herschel space telescope.

Optical Image

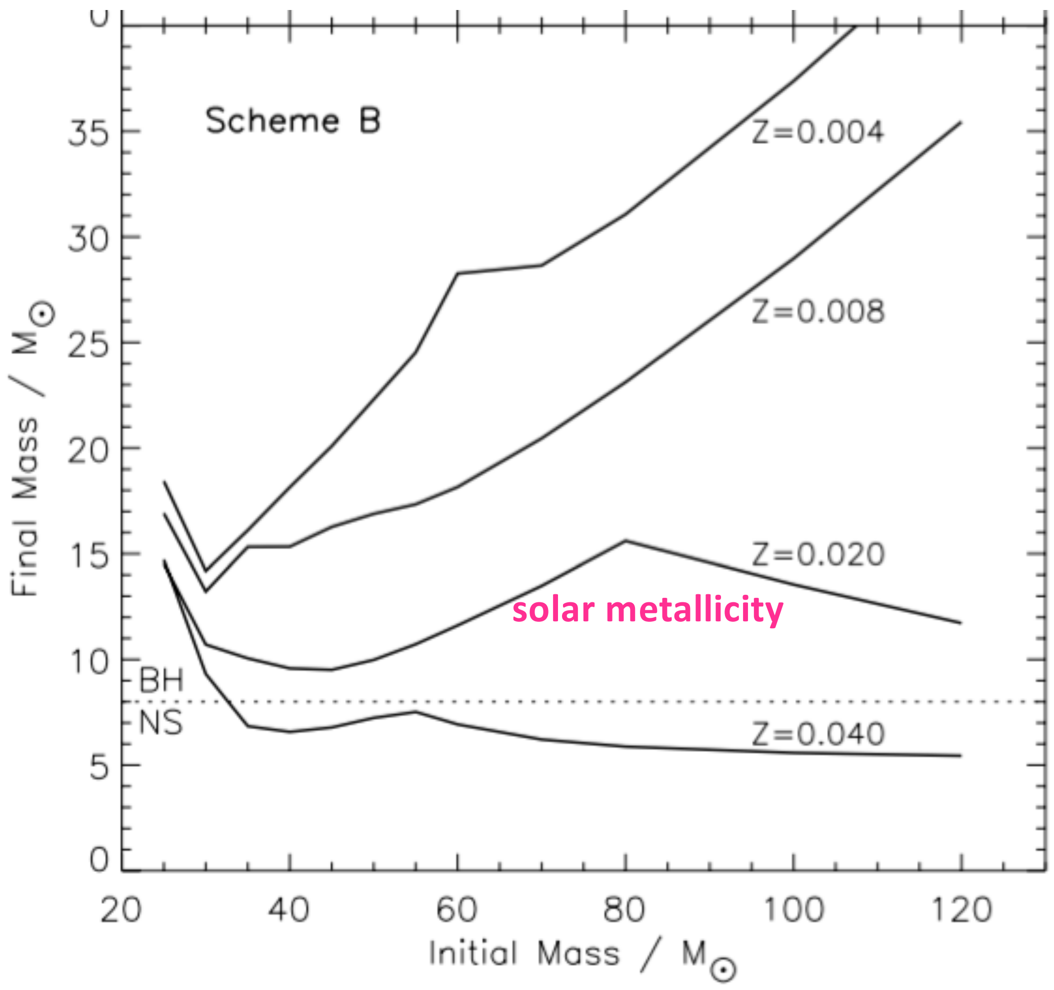
V838 Mon: Another example of mass-loss from red-supergiant star.

A ***Wolf-Rayet Star*** (WR star). It is a naked helium star that has lost its hydrogen envelope. A star of this kind has usually very strong winds.



Eta Carinae nebula: The star at the center has a mass of about 100 Msun. This star underwent a great eruption in the 1840s, ejecting about 10 Msun.

Final mass of single stars as a function of metallicity



Eldridge & Vink 2006

III. Conditions for neutron star and black hole formation

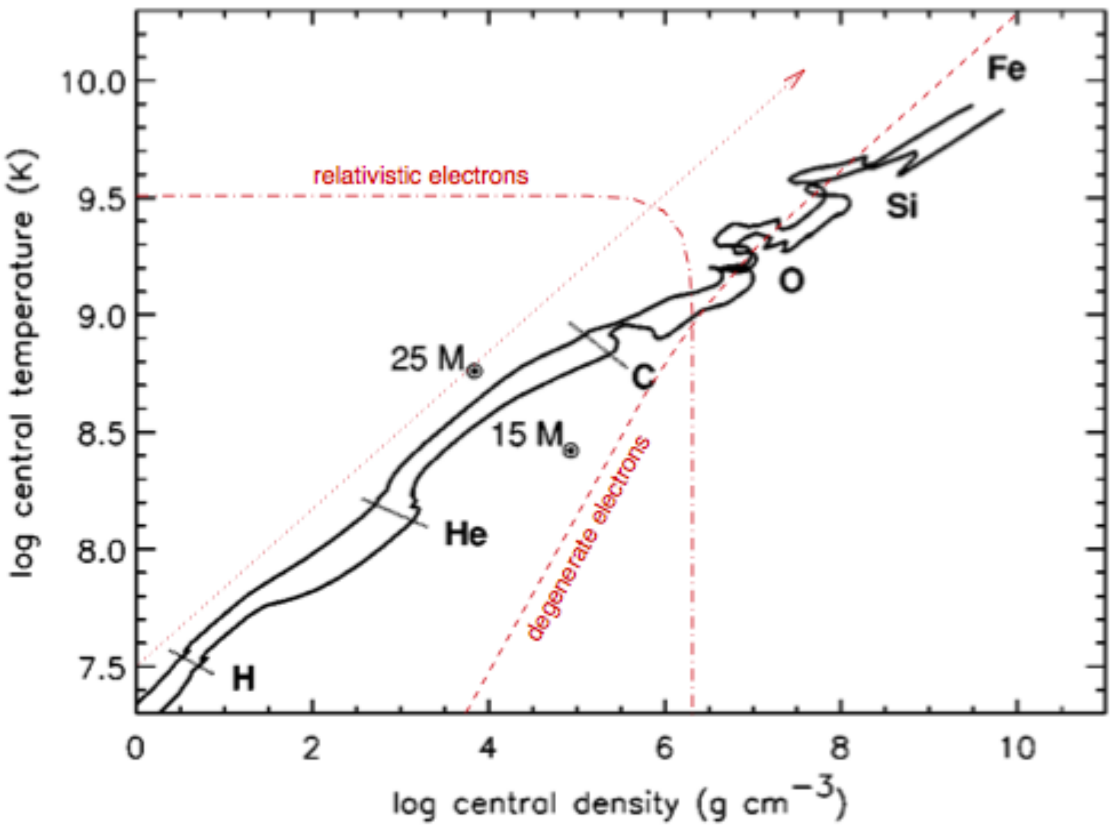


Figure 12.5. Evolution of central temperature and density of 15 $M_\odot$ and 25 $M_\odot$ stars at $Z = 0.02$ through all nuclear burning stages up to iron-core collapse. The dashed line indicated where electrons become degenerate, and the dash-dotted line shows where electrons become relativistic ($\epsilon_e \approx m_e c^2$). The dotted line and arrow indicates the trend $T_c \propto {\rho_c}^{1/3}$ that is expected from homologous contraction. Non-monotonic (non-homologous) behaviour is seen whenever nuclear fuels are ignited and a convective core is formed. Figure adapted from Woosley, Heger & Weaver (2002, Rev. Mod. Ph. 74, 1015).

The stellar evolution at the final stages is highly non-linear with complicated convection history, and the final stellar structure is not easy to predict robustly.

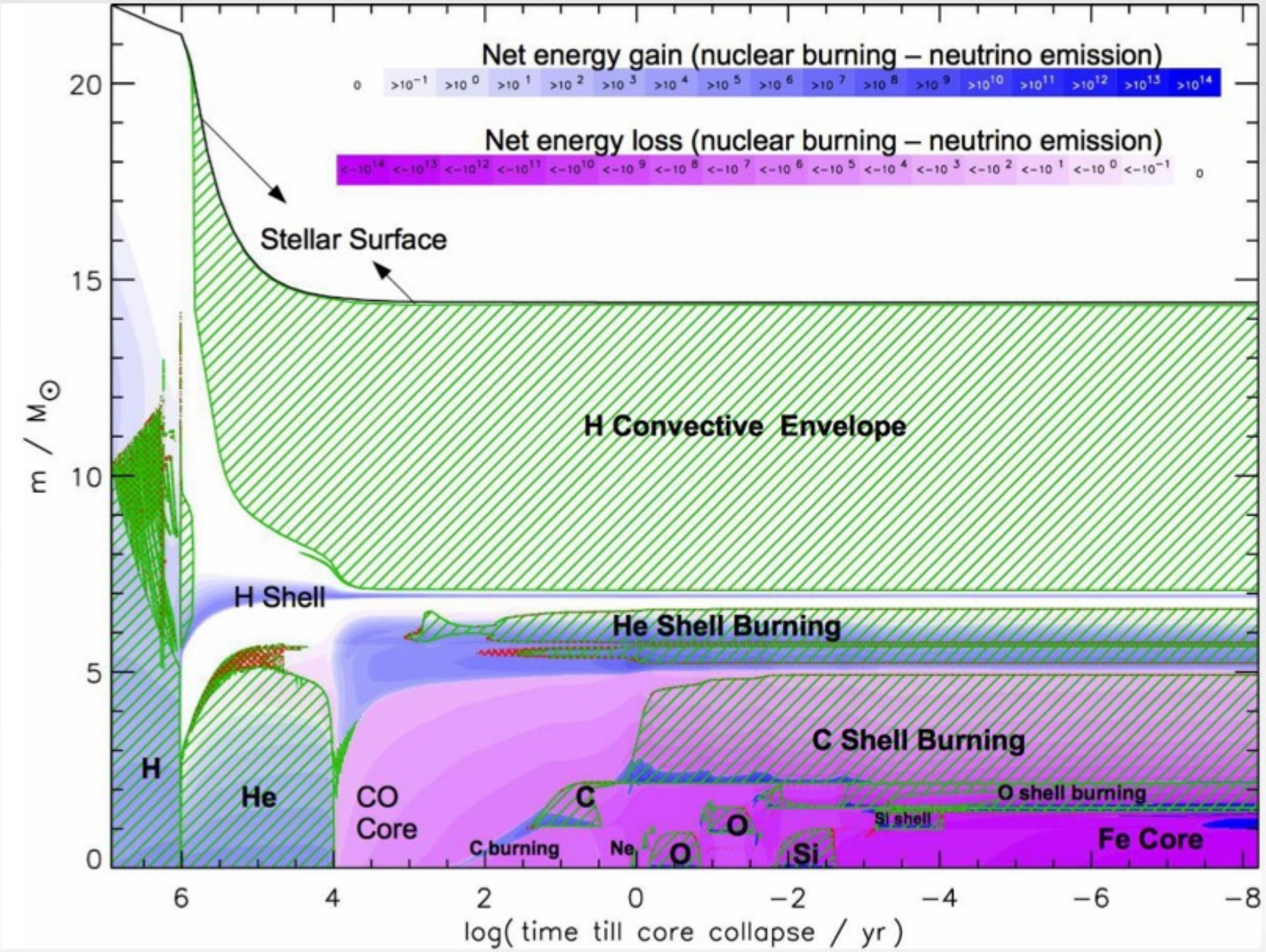
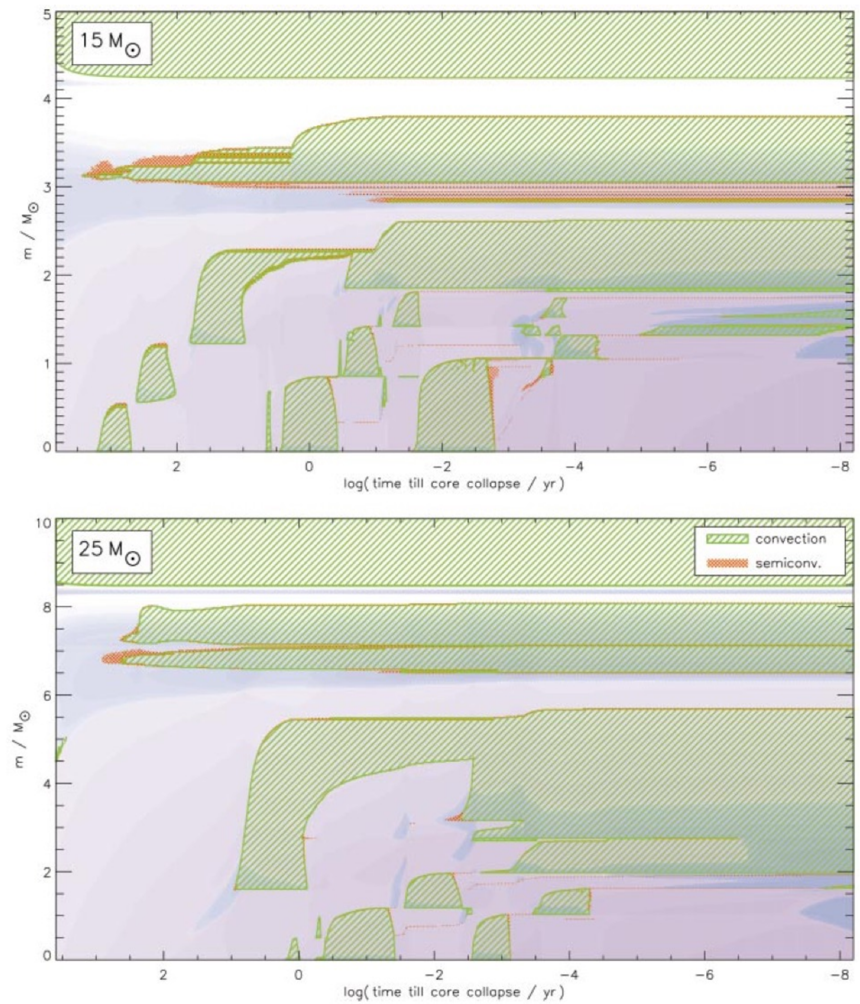


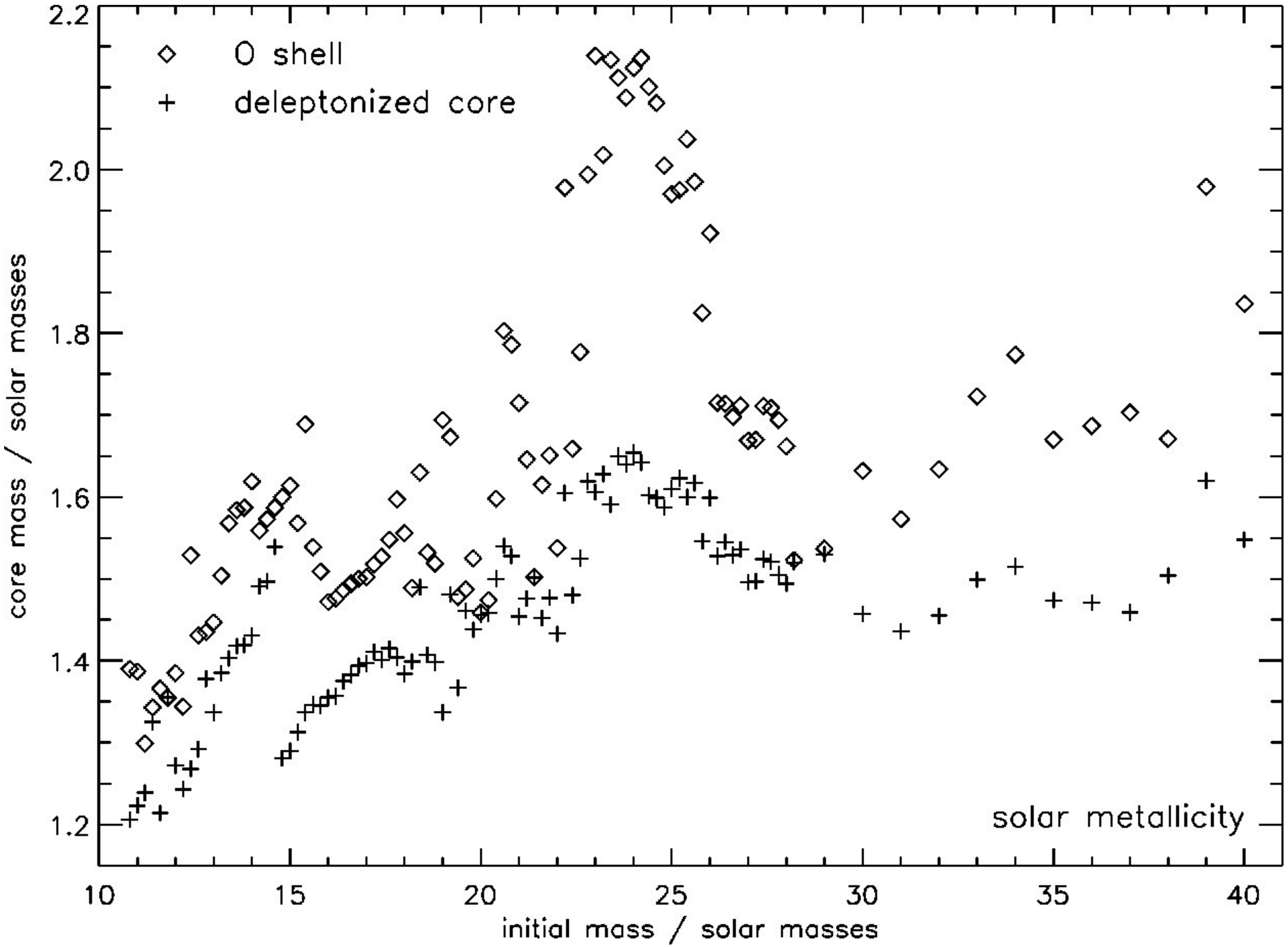
Figure from A. Heger

Pre-SN Evolution



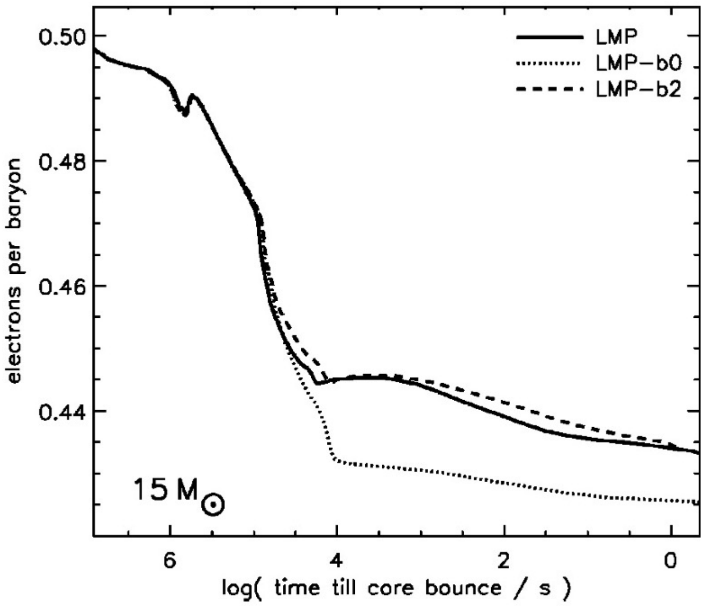
Woosley & Heger 2002

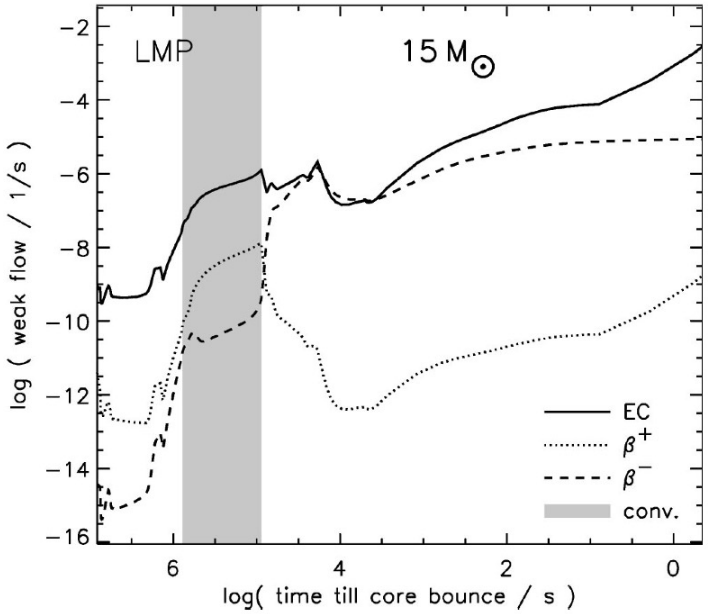
Due to the complex history with convection and shell burning, the iron core size is not a monotonic function of the intial mass.



Woosley et al. 02

Pre-SN Evolution





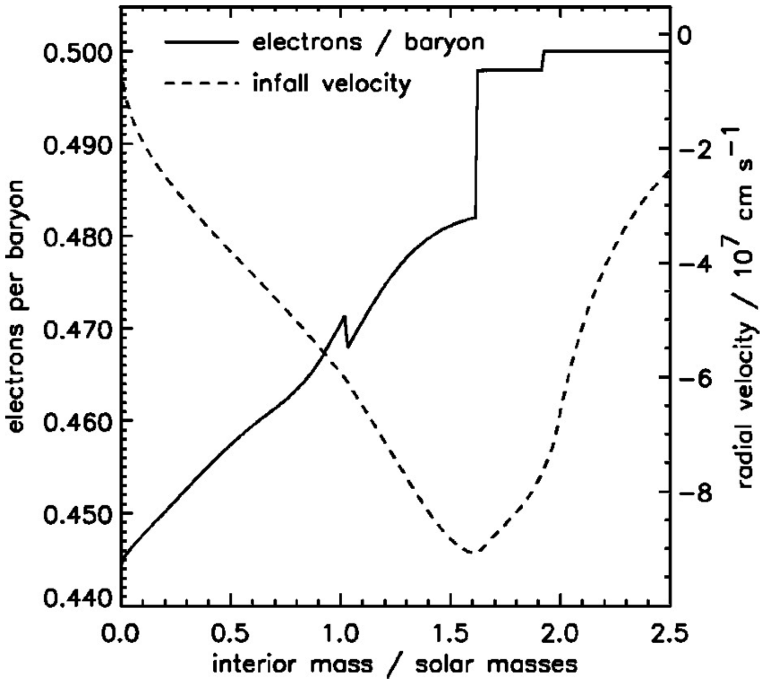
Woosley & Heger 2002

Due to the electron capture, the number of electron per baryon (Y_e) decreases:

$$Y_e = \sum \frac{Z_i X_i}{A_i} \tag{1}$$

Beta-decays also affect the final Y_e.

Pre-SN Evolution



Woosley & Heger 2002

The Chandrasekhar mass:

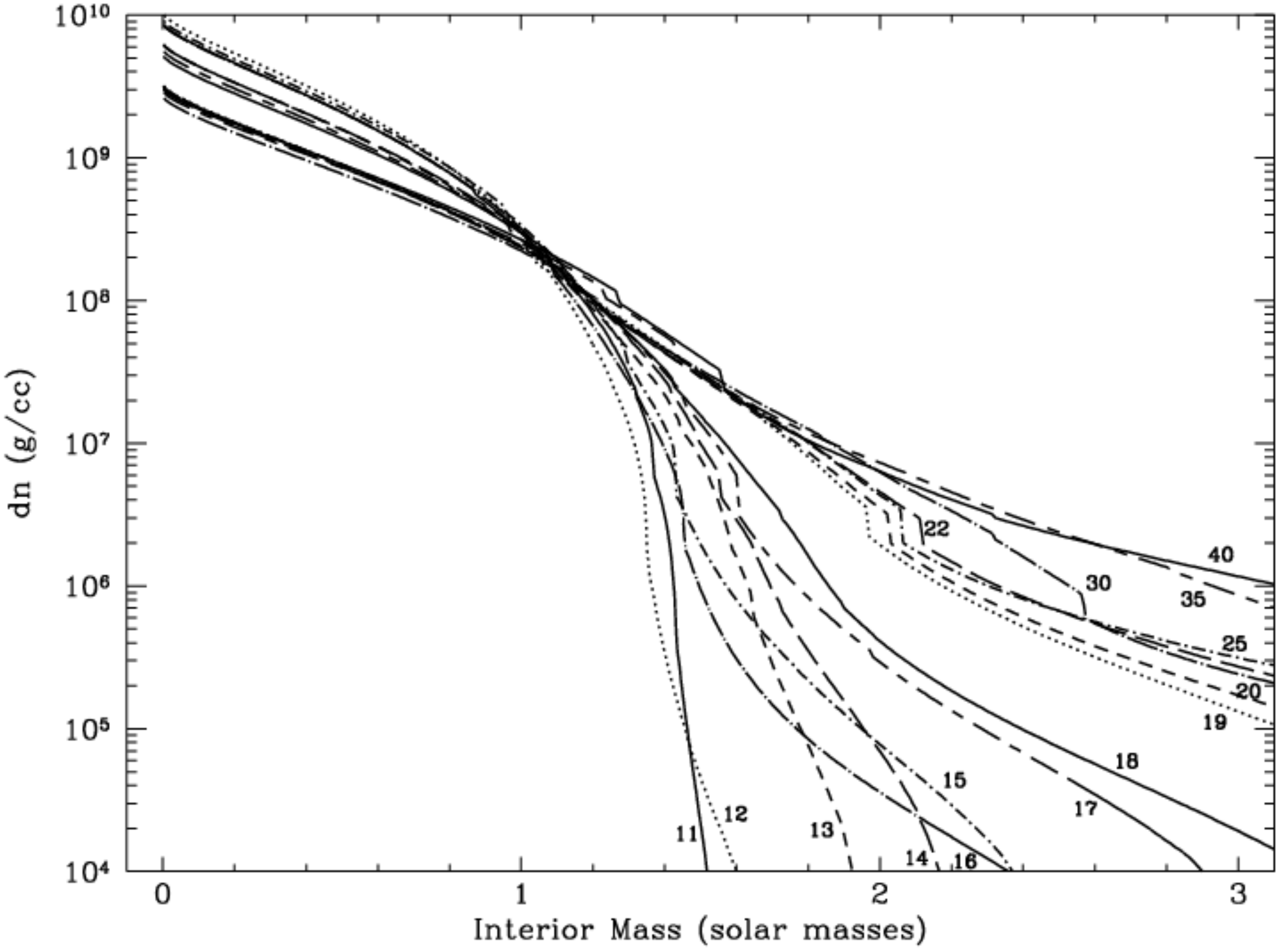
$$M_{\rm Ch} = 5.83 Y_e^2 \quad . \qquad (2)$$

For $Y_e = 0.45$, we have $M_{\rm Ch} = 1.18\ {\rm M}_\odot$. The effective Chandrasekhar mass:

$$M_{\rm Ch,eff} = M_{\rm Ch}\left[1 + \left(\frac{\pi^2 k^2 T^2}{\epsilon_{\rm F}^2}\right)\right] \qquad (3)$$

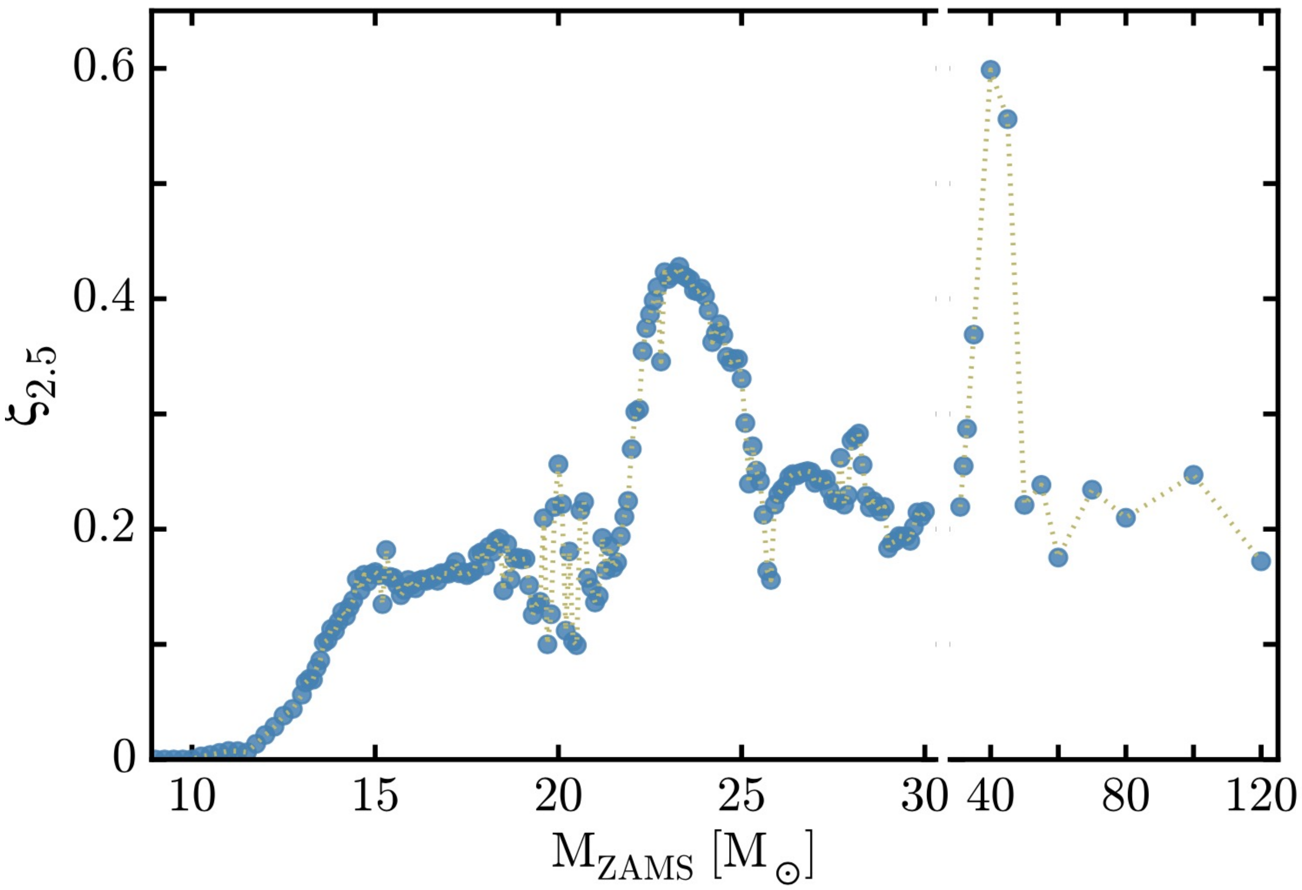
where $\epsilon_{\rm F} = 1.11(\rho_7 Y_e)^{1/3}$ MeV. For an iron core of 15 ${\rm M}_\odot$ star, we have roughly $M_{\rm Ch,eff} \simeq 1.34\ {\rm M}_\odot$.

binding energy becomes higher for a higher initial mass.

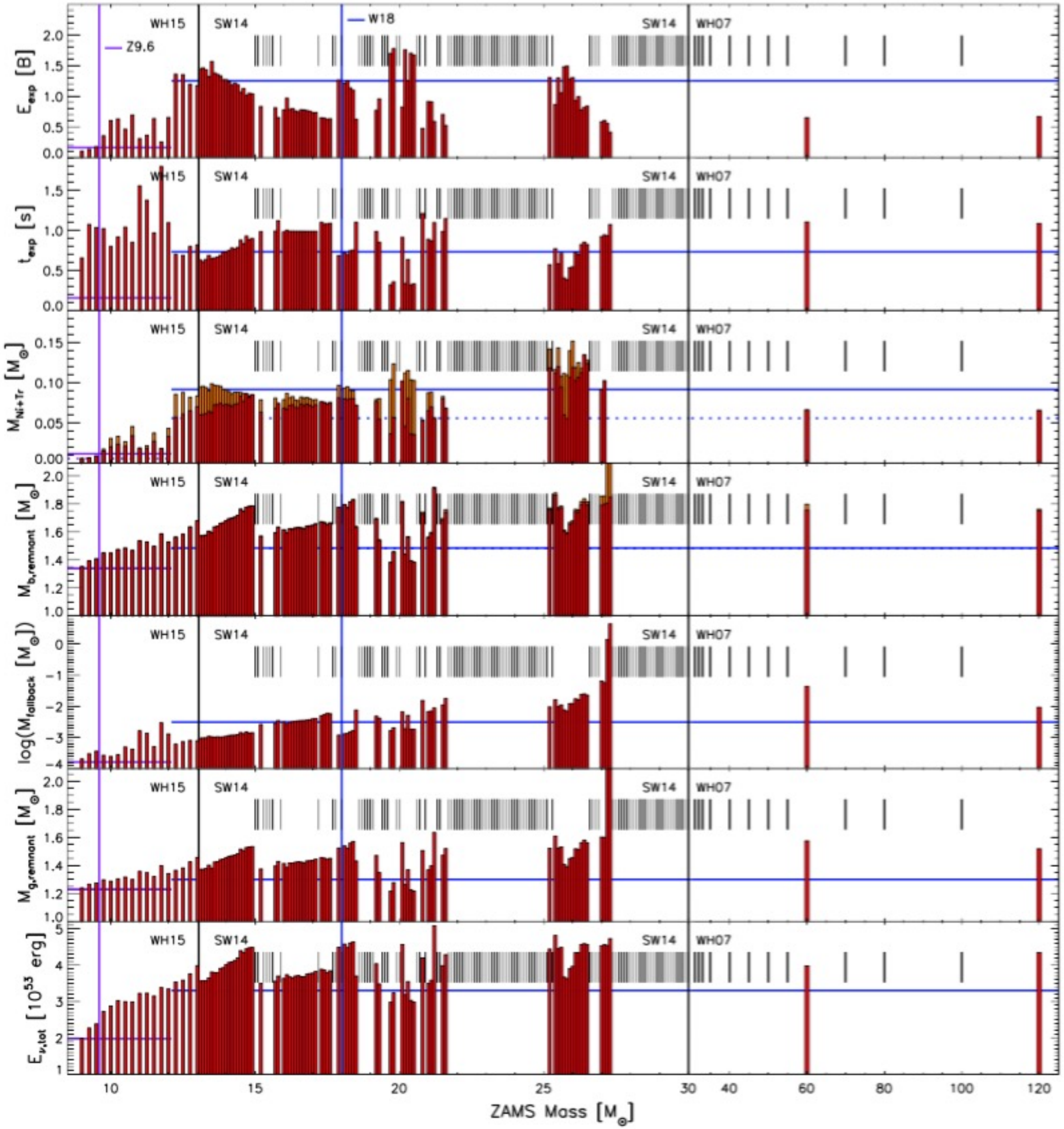


Compactness parameter

$$\xi_M = \frac{M/\mathrm{M}_\odot}{R(M)/1000\,\mathrm{km}}\bigg|_{t_{\mathrm{bounce}}}$$



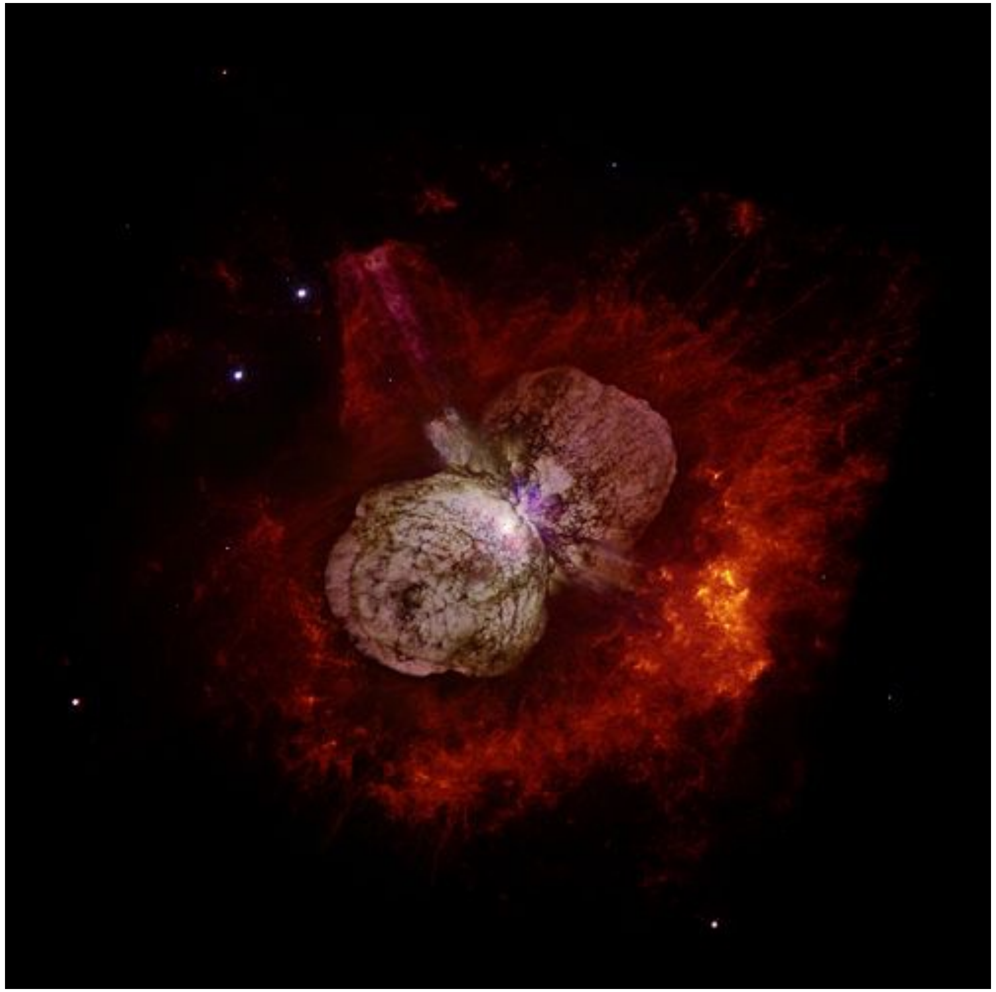
Sukhbold et al. 2016



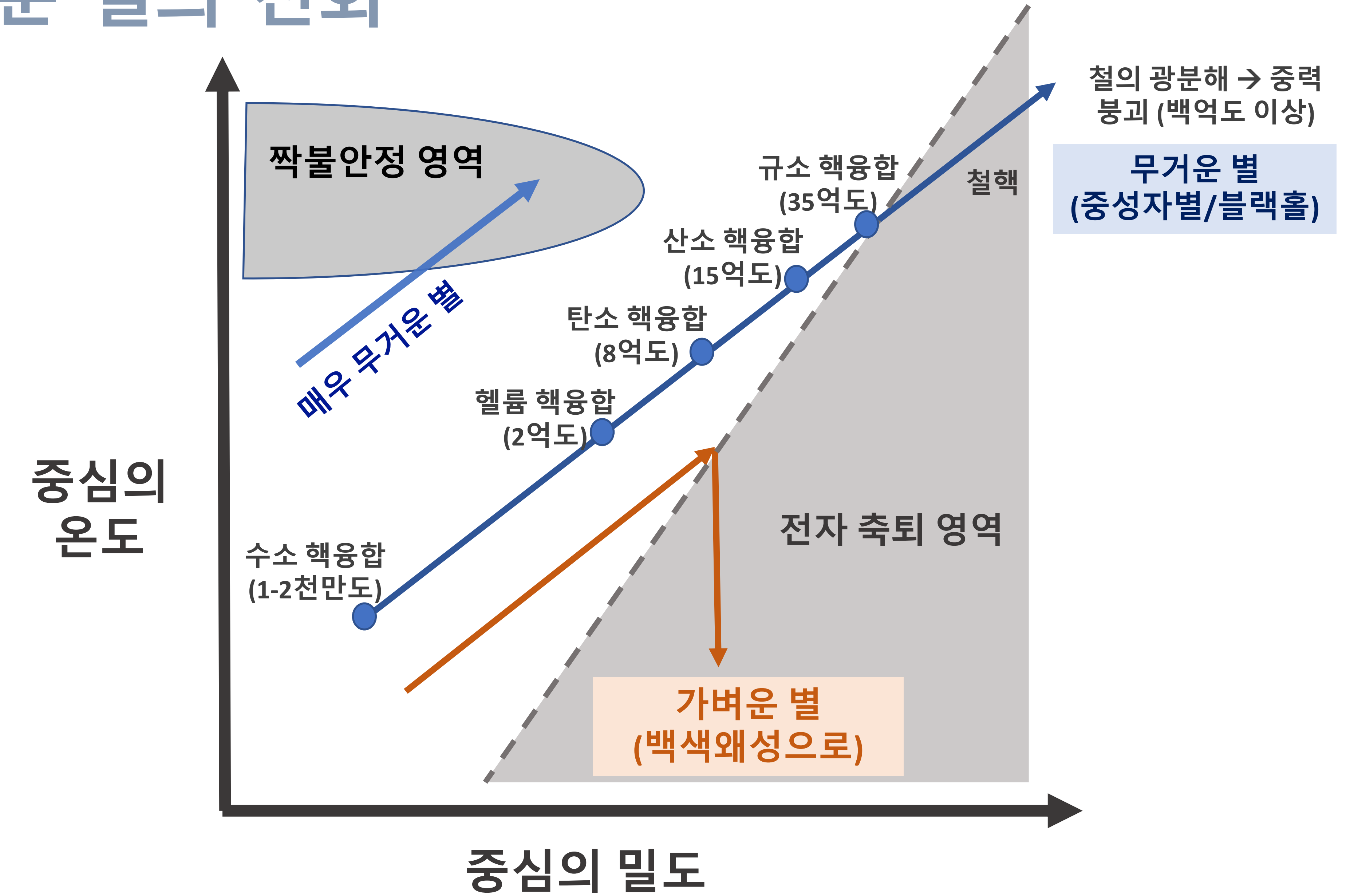
Sukhbold et al. 2016

가장 무거운 별은?

- **R136a1**: 265 M_{sun} (대마젤란은하)
- **R136a2**: 195 M_{sun} (대마젤란은하)
- **VFTS 682** : 150 M_{sun} (대마젤란은하)
- **R136a3** : 135 M_{sun} (대마젤란은하)
- **NCG 3603-B**: 132 M_{sun} (우리은하)
- **Eta Carinae A**: 120 M_{sun} (우리은하)



매우 무거운 별의 진화

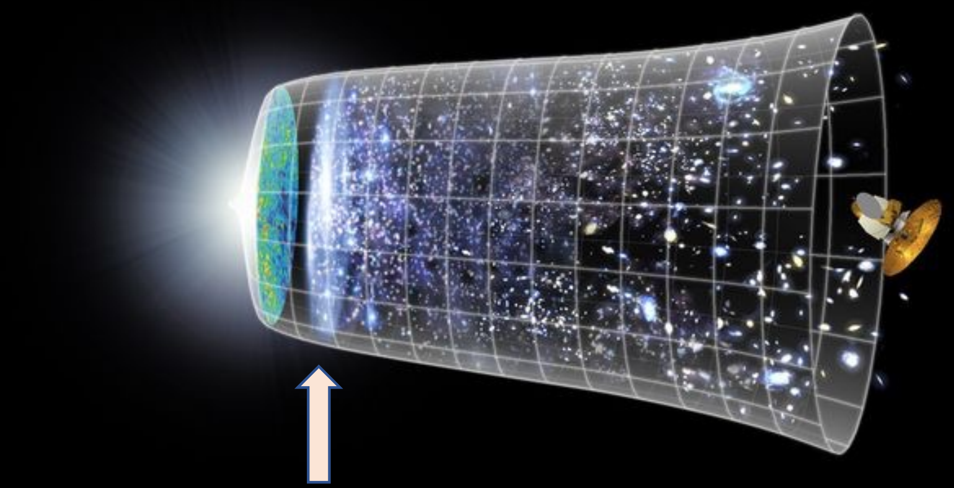


중심의 온도
중심의 밀도
짝불안정 영역
매우 무거운 별
수소 핵융합 (1-2천만도)
헬륨 핵융합 (2억도)
탄소 핵융합 (8억도)
산소 핵융합 (15억도)
규소 핵융합 (35억도)
철핵
철의 광분해 → 중력 붕괴 (백억도 이상)
무거운 별 (중성자별/블랙홀)
전자 축퇴 영역
가벼운 별 (백색왜성으로)

매우 무거운 별의 폭발

태양 질량의 140배에서 300배 :
짝불안정 초신성

(pair instability supernova)



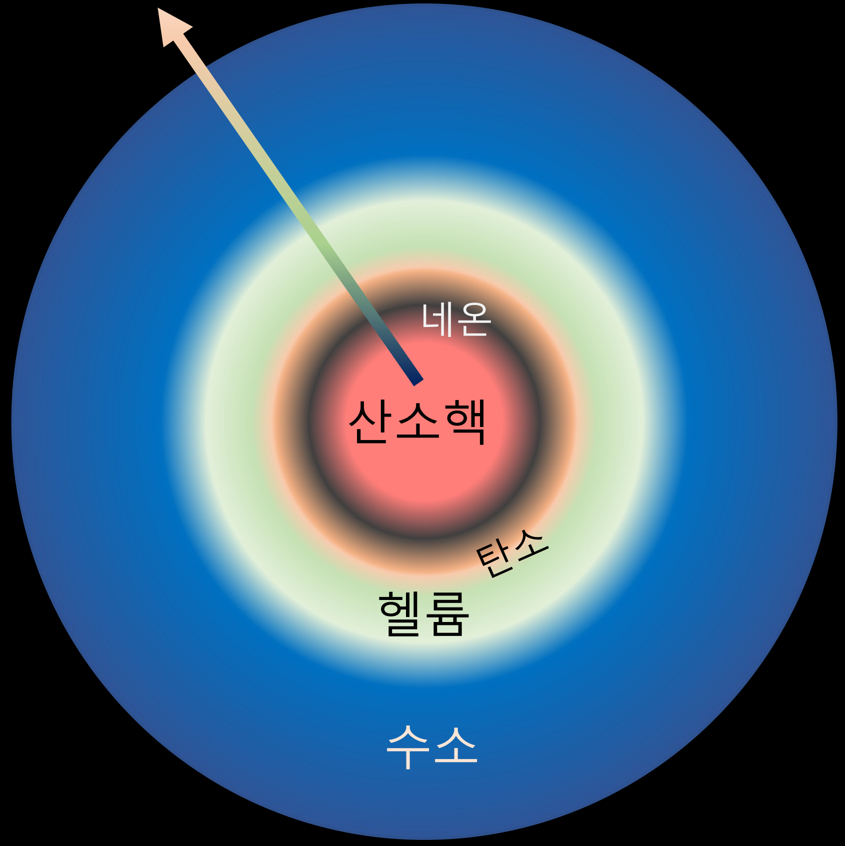
첫번째 별들

산소핵의 높은 온도:
감마선에 의한 짝 생성

$$\gamma \rightarrow e^- + e^+$$

복사압 감소
에 따른
중력붕괴

폭발적 산소
핵반응에 따른
별의 폭발



NASA/WMAP Science Team

Pair-Instability Gap for Black Hole