Lecture 15

Explosive Nucleosynthesis and the r-Process

As the shock wave passes through the star, matter is briefly heated to temperatures far above what it would have experienced had it burned in hydrostatic equilibrium. This material expands, then cools nearly adiabatically. The time scale for the cooling is approximately the *hydrodynamic time scale*, though a little shorter (because speeds are faster than free fall).

For (post-helium) burning in *hydrostatic equilibrium*, recall we had

$$\boxed{\langle \varepsilon_{\rm nuc} \rangle \approx \langle \varepsilon_\nu \rangle}$$

hydrostatic nucleosynthesis
advanced stages of stellar evolution

For exp*losive nucleosynthesis we have instead* :

$$\boxed{\tau_{\rm nuc}(T_{shock}) \le \tau_{HD}}$$

$$\rho(t) = \rho_{\rm shock} \exp(-t/\tau_{HD})$$

$$T(t) = T_{\rm shock} \exp(-t/3\tau_{HD})$$

$$\boxed{\tau_{HD} = \frac{446 \text{ sec}}{\sqrt{\rho_{shock}}}}$$

$$\rho \propto T^3$$

$$T_{shock} \approx 2.4\, x 10^9\, R_9^{-3/4}$$

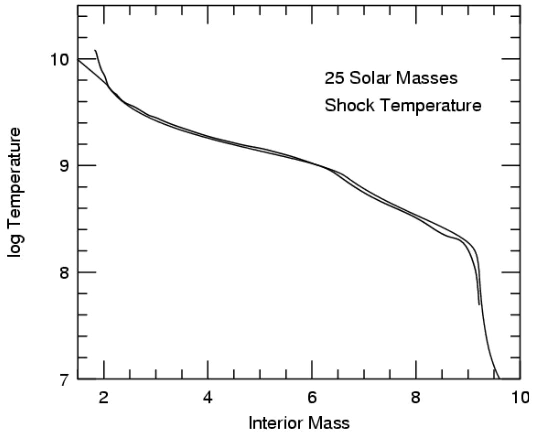
$$\rho_{shock} \approx 2\ \rho_{\rm o}$$

Except near the "mass cut", the shock temperature to which the explosive nucleosynthesis is most sensitive is given very well by

$$\boxed{\frac{4}{3}\pi R^3 a T^4 \approx \text{Explosion energy}}$$

$$\approx 10^{51} \text{ erg}$$

This is because of the near constancy of pressure behind the shock and the dominance of radiation



Example:

Any carbon present inside of 10^9 cm will burn explosively since:

At $T_9 = 2$ $\quad \lambda_{12,12} \approx 4.3 \times 10^{-4} \left(\frac{T_9}{2}\right)^{20}$

$$\frac{dY_{12}}{dt} = -2\, Y_{12}^2\, \rho\, \lambda_{12,12}/2$$

$$\tau_{12,12} = (\rho Y_{12} \lambda_{12,12})^{-1} = \frac{12}{\rho X_{12} \lambda_{12,12}}$$

$$= 1.9 \left(\frac{2}{T_9}\right)^{20} = 45/\sqrt{10^5} = 0.14 \text{ s}$$

0.1 τ_{HD}

$X_{12} \approx 0.15$

$\rho \sim 10^5$

$$\Rightarrow \underline{\underline{T_9 \simeq 2.3}}$$

Where in the star does this occur ?

$$10^{51} = \frac{4}{3}\pi R^3 a (2.3 \times 10^9)^4 \Rightarrow \underline{\underline{R = 1.0 \times 10^9 \text{ cm}}}$$

<u>nb.</u> $T_p \propto R^{-3/4}$

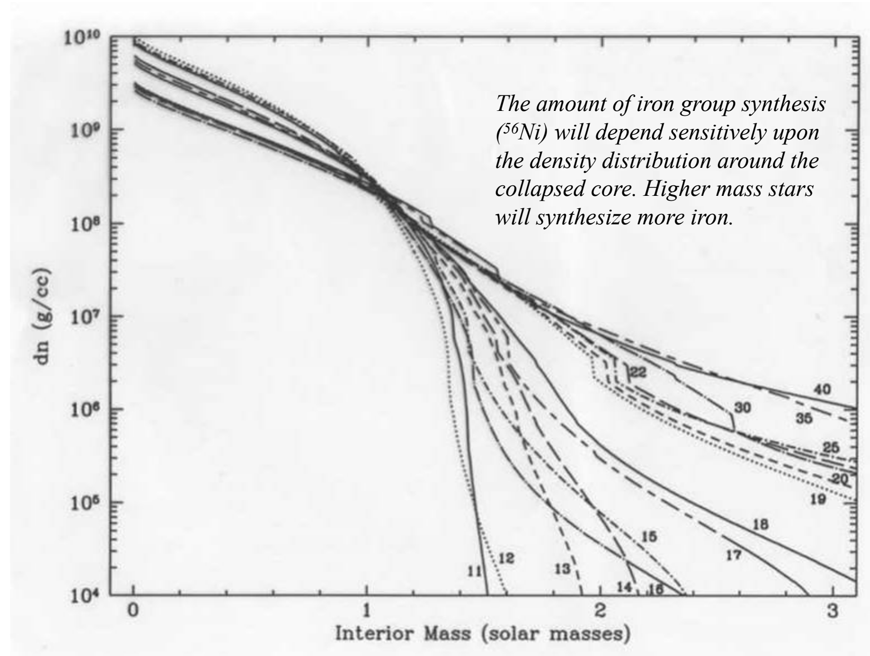
about 3.2 $M_\odot$ in the 25 $M_\odot$ presN star

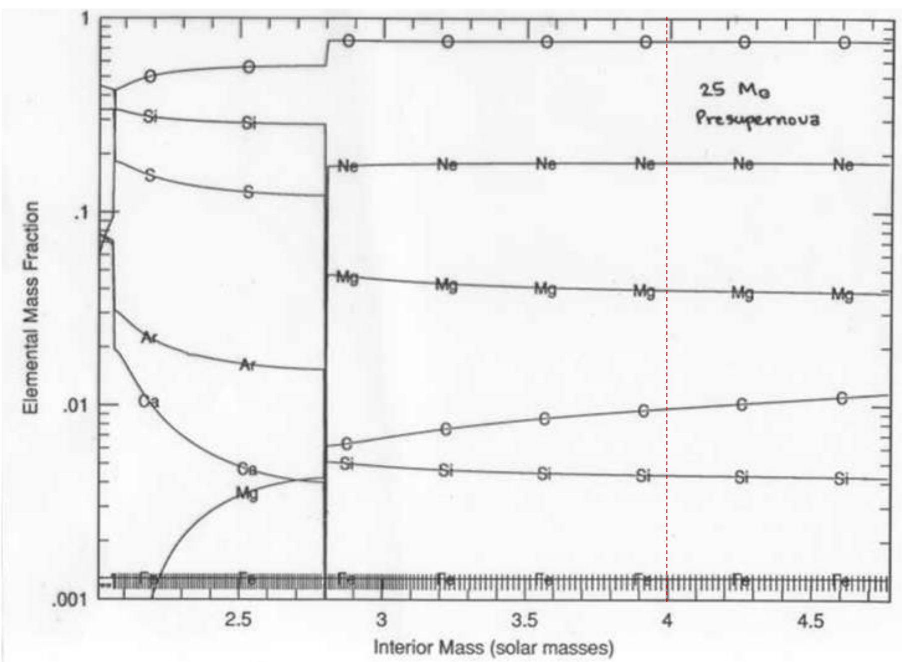
Conditions for explosive burning in one 25 $M_{\odot}$ model (E_0 = 1.2 B):

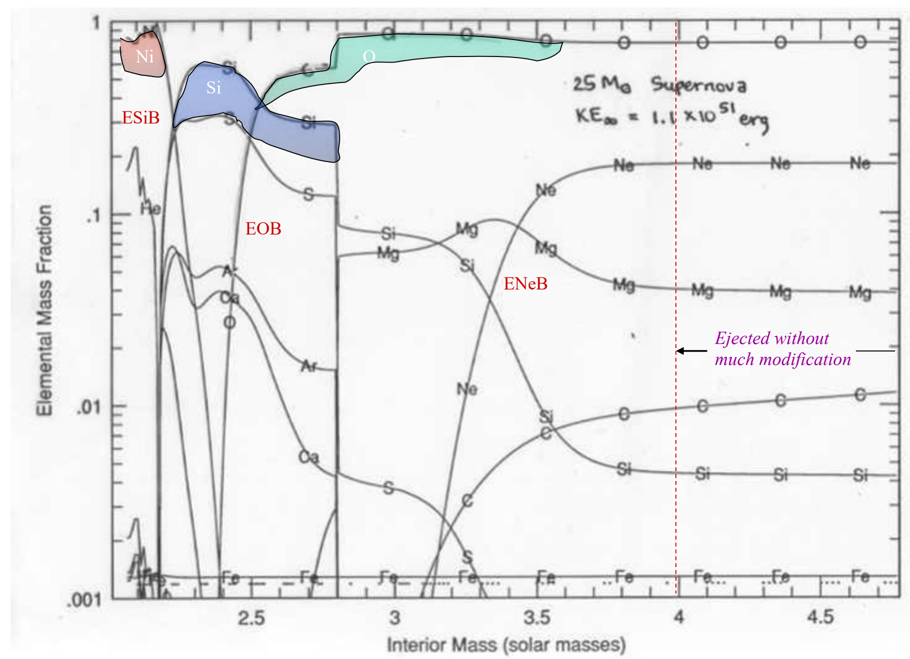
To nse	$T_9 > 5$	$R_9 < 0.38$	*yes*
Silicon	4 – 5	0.38 – 0.51	*yes*
Oxygen	3 – 4	0.51 – 0.74	*yes*
Neon	2.5 – 3	0.74 – 1.0	*a little*
Carbon	2.0 – 2.5	1.0 – 1.3	*no*
Helium	>0.5	< 8	*no*
Hydrogen	>0.2	< 27	*no*

$$R_9 = \left(\frac{T_9}{2.4}\right)^{-4/3}$$

Roughly speaking, everything that is ejected from inside 3800 km in the presupernova star will come out as iron-group elements.







Produced pre-explosively and just ejected in the supernova:

- Helium
- Carbon, nitrogen, oxygen
- The s-process
- Most species lighter than silicon

Produced in the explosion:

- Iron and most of the iron group elements – Ti, V, Cr, Mn, Fe Co, Ni
- The r-process (?)
- The neutrino process – F, B

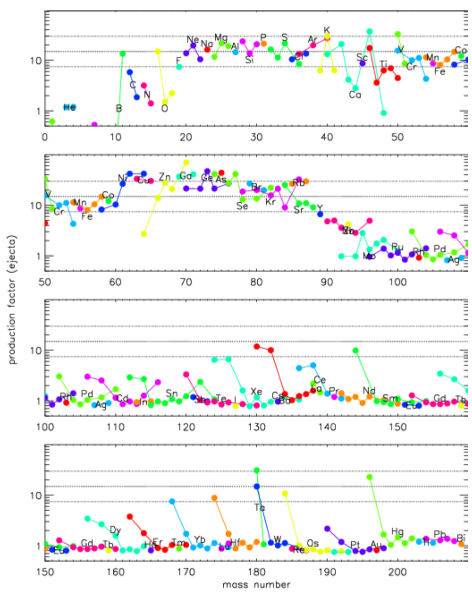
Produced both before and during the explosion:

- The intermediate mass elements – Si, S, Ar, Ca
- The p-process (in oxygen burning and explosive Ne burning)

Explosive Nucleosynthesis

Fuel	Main Product	Secondary Products	Temp (10^9 K)	Time (sec)
Innermost Ejecta	r- process	-	>10 low Y_e	about 1
Si, O	^{56}Ni	Iron group	> 4	0.1
O	Si,S	Cl, Ar K, Ca	3 - 4	1
O, Ne	O, Mg Ne	Na, Al P	2 - 3	5
		p - process ^{11}B, ^{19}F	"	"

25 Solar Masses; Rauscher et al. (2002), parameterized explosion
KE = 1.2 x 10^{51} erg at infinity (requires strong successful explosion)



The nucleosynthesis that results from explosive silicon burning is sensitive to the density (and time scale) of the explosion.

1) High density (or low entropy) NSE, and long time scale:

Either the material is never photodisintegrated even partially to α–particles or else the α–particles have time to reassemble into iron-group nuclei. The critical (slowest) reaction rate governing the reassembly is $\alpha(2\alpha,\gamma)^{12}$C which occurs at a rate proportional to ρ^2.

If as T $\rightarrow 0$, X_n, X_p, and $X_\alpha \rightarrow 0$ then one gets pretty much the unmodified "normal" results of nuclear statistical equilibrium calculated e.g.,at T_9 = 3 (fairly indedendent of ρ).

Abundant nuclei at $\eta = 0.002 - 0.004$

^{56,57}Ni, ^{55}Co, ^{52,53,54}Fe, ^{48,49,50}Cr, ^{51}Mn

Products after all decays are complete:

^{54,56,57}Fe, ^{55}Mn, ^{48,49}Ti, ^{50,51,52,53}Cr, ^{51}V

2) Low density or rapid expansion → the "α-rich" freeze out

If all the α's from photodisintegration cannot reassemble on τ_{HD}, then the composition will be moditied at late times by α-capture. The composition will "freeze out" with free α-particles still present (and, in extreme cases, free n's or p's). The NSE composition at low T willbe modified by reactions like

$$^{54}Fe(\alpha,\gamma)^{58}Ni \qquad ^{56}Ni(\alpha,p)^{59}Cu$$
$$^{56}Ni(\alpha,\gamma)^{60}Zn \qquad ^{40}Ca(\alpha,\gamma)^{44}Ti$$
$$^{57}Ni(\alpha,\gamma)^{61}Zn \qquad ^{58}Ni(\alpha,\gamma)^{62}Zn \qquad etc.$$

Abundant:

^{44}Ti, ^{56,57,58}Ni, ^{59}Cu, ^{60,61,62}Zn,(^{64,66}Ge)

Produced :

^{44}Ca, ^{56,57}Fe, ^{58,60,61,62}Ni, ^{59}Co,(^{64,66}Zn)

Both kinds of freeze-out occur in a typical explosion

3) Explosive oxygen burning $(3 \le T_9 \le 4)$

Makes pretty much the same products as ordinary oxygen burning ($T_9 \approx 2$) at low $\eta \approx 0.002$ ($Z/Z_\odot$)

Principal Products:

^{28}Si, ^{32,33,34}S, ^{35,37}Cl, ^{36,38}Ar
^{39,41}K, ^{40,42}Ca, ^{46}Ti, ^{50}Cr

4) Explosive neon and carbon burning $(2 \le T_9 \le 3)$

Same products as stable hydrostatic buring

More ^{26}Al, the p-process or γ-process.

5) Explosive H and explosive He burning.

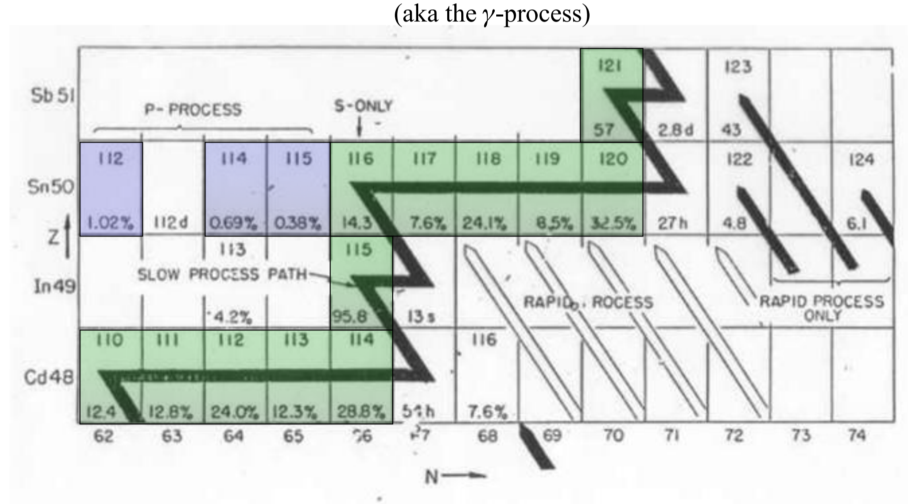
The former occurs in novae; the latter in some varieties of Type Ia supernovae. Discuss later.

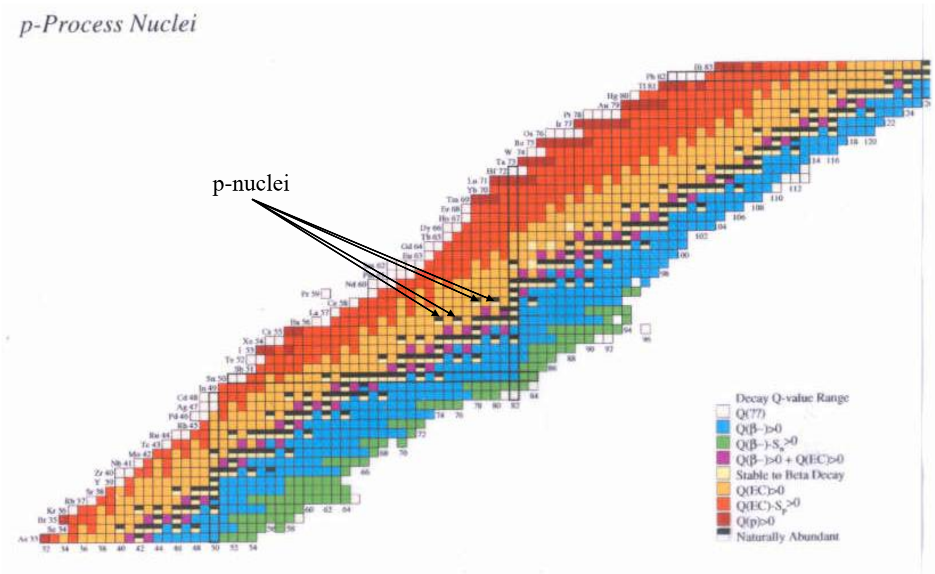
The "p -" or

γ - Process

At temperatures ~2 x 10^9 K before the explosion (oxygen burning) or between 2 and 3.2 x 10^9 K during the explosion (explosive neon and oxygen burning) partial photodisintegration of pre-existing s-process seed makes the proton-rich elements above the iron group.

The *p*-Process
(aka the γ-process)





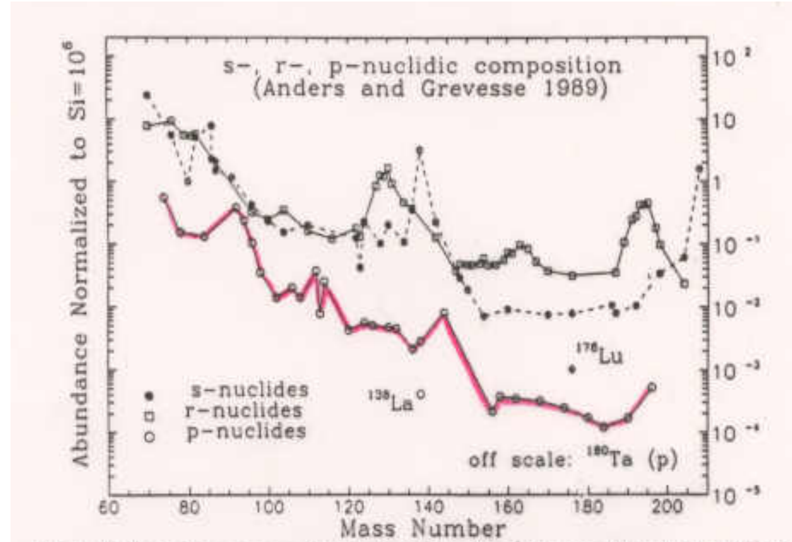
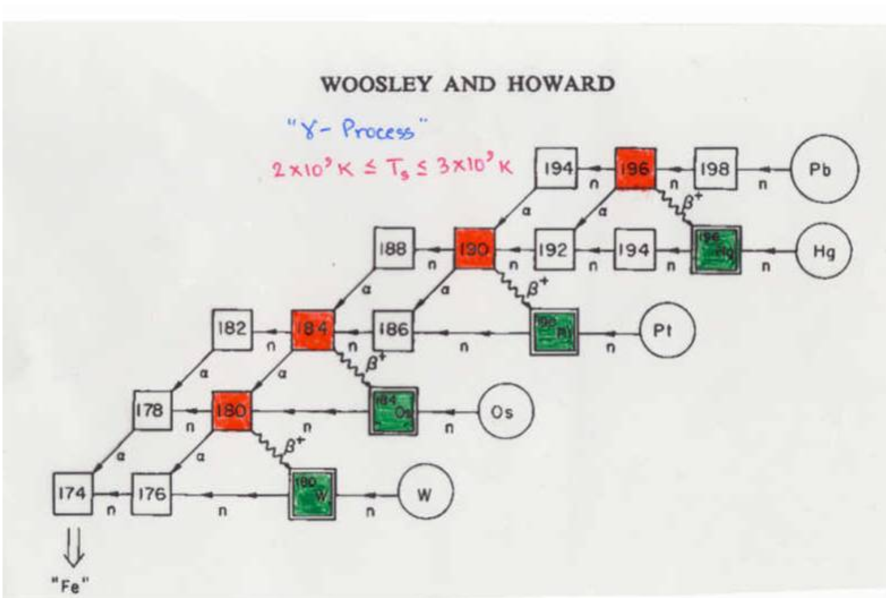
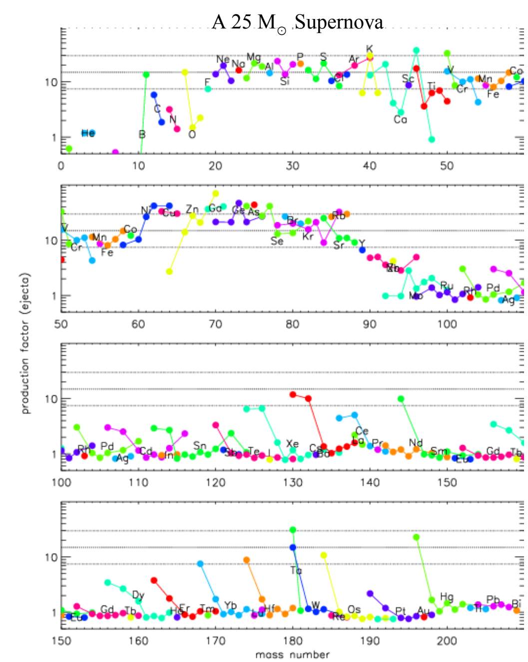


Fig. 1. Abundances of p, r, and s-nuclides in the SAD (Anders and Grevesse 1989). Note that ^{180}Ta has the very low abundance of 2.5×10^{-6}. (From Arnould 1991)





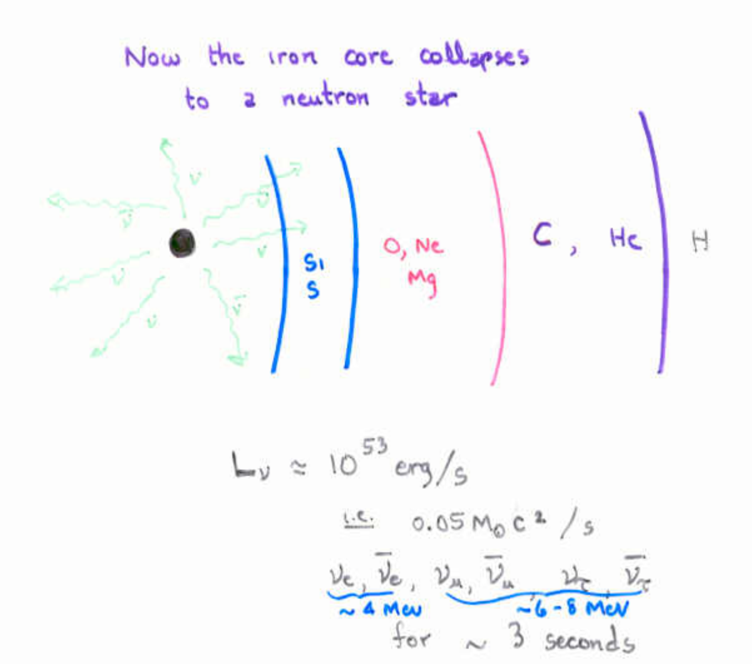
Problems below A ~ 130.

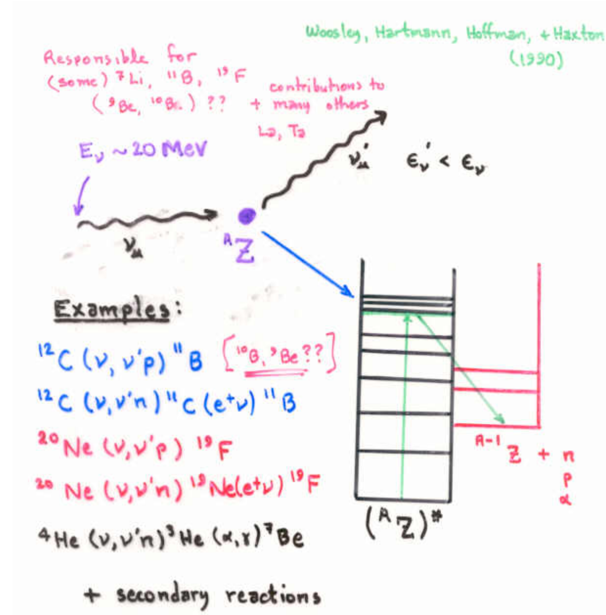
Summary: γ−Process

- Makes nuclei traditionally attributed to the "p-process" by photodisintegration of pre-existing s-process nuclei. The abundance of these seeds is enhanced – at least for A < 90 – by the s-process that went on in He and C burning.
- Partially produced in oxygen shell burning before the collapse of the iron core, but mostly made explosively in the carbon, neon, and oxygen-rich shells that experience shock temperatures between 2 and 3.2 billion K.
- Production factor ~100 in about 1 solar mass of ejecta. Enough to make solar abundances
- A secondary (or tertiary) process. Yield is proportional to abundance of s-process in the star.
- There remain problems in producing sufficient quantities of p-nuclei with atomic masses between about 90 and 120, especially ^{92}Mo.

The Neutrino Process (ν–process)

The neutrino flux from neutron star formation in the center can induce nuclear transmutation in the overlying layers of ejecta. The reactions chiefly involve μ and τ-neutrinos and neutral current interactions. Notable products are ^{11}B, ^{19}F. ^{138}La, ^{180}Ta, and some ^{7}Li and ^{26}Al.





Production factor relative to solar normalized to ^{16}O production as a function of μ and τ neutrino temperature (neutral current) and using 4 MeV for the electron (anti-)neutrinos (for charged current only). 6 MeV is now considered a more likely value for $T_{\mu\tau}$

Product	$15\ M_\odot$				$25\ M_\odot$			
	6 MeV		**8 MeV**		**6 MeV**		**8 MeV**	
	WW95	**This work**	**WW95**	**This work**	**WW95**	**This work**	**WW95**	**This work**
^{11}B	1.65	1.88	3.26	3.99	0.95	1.18	1.36	1.85
^{19}F	0.83	0.60	1.28	0.80	0.56	0.32	1.03	0.53
^{15}N	0.46	0.49	0.54	0.58	0.09	0.12	0.15	0.19
^{138}La		0.97		1.10		0.90		1.03
^{180}Ta		2.75		3.07		4.24		5.25

*Heger et al,, 2005, Phys Lettr B, **606**, 258*

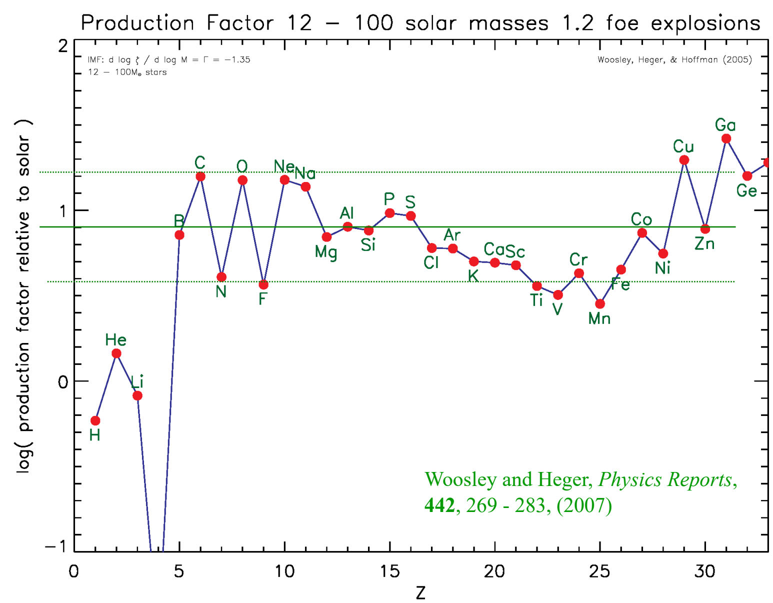
Integrated Ejecta

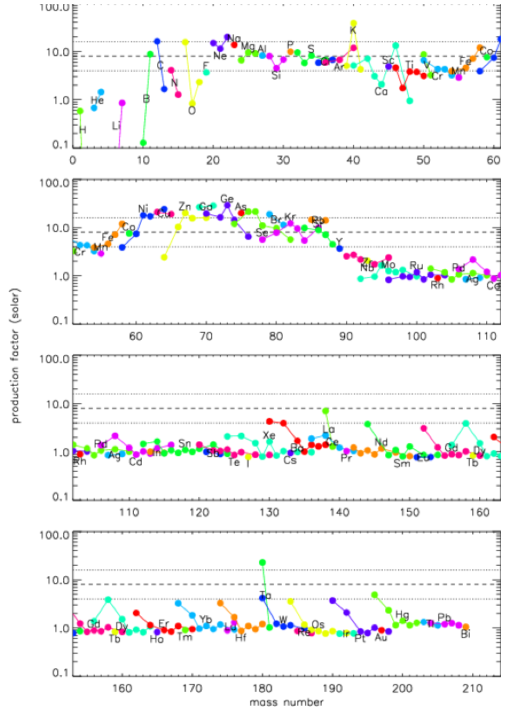
Averaged yields of many supernovae integrated either over an IMF or a model for galactic chemical evolution.

Survey - Solar metallicity:

(Woosley and Heger 2007)

- *Composition – Lodders (2003); Asplund, Grevesse, & Sauval (2004)*
- *32 stars of mass 12, 13, 14, 15, 16, 17, 18, 19, 20, 21, 22, 23, 24, 25, 26, 27, 28, 29, 30, 31, 32, 33, 35, 40, 45, 50, 55, 60, 70, 80, 100, 120 solar masses.*
- *Evolved from main sequence through explosion with two choices of mass cut ($S/N_AkT = 4$ and Fe-core) and two explosion energies (1.2 B, 2.4 B) – 128 supernova models*
- *Averaged over Salpeter IMF*





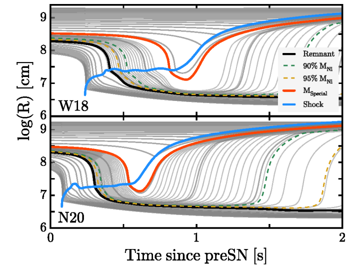
Isotopic yields for 31 stars averaged over a Salpeter IMF, $\Gamma = -1.35$

Intermediate mass elements (23< A < 60) and s-process (A = 60 – 90) well produced.

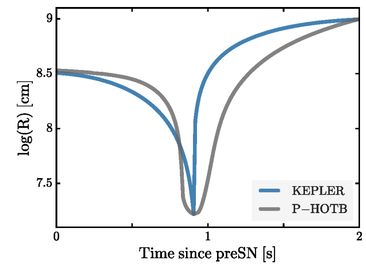
Carbon and Oxygen over-produced.

p-process deficient by a factor of ~4 for A > 130 and absent for A < 130

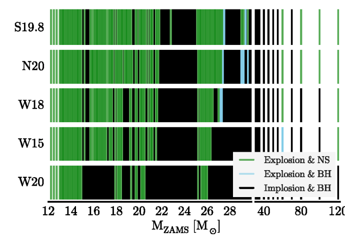
Sukhbold et al (2016) – 300 supernova models – 2000 nuclei



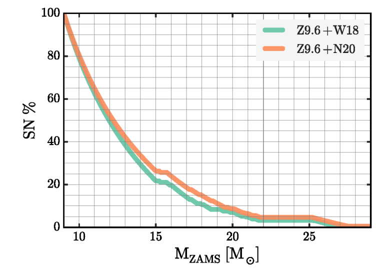
Ertl and Janka (P-HOTB)



Mapped into KEPLER



Successful explosions in green

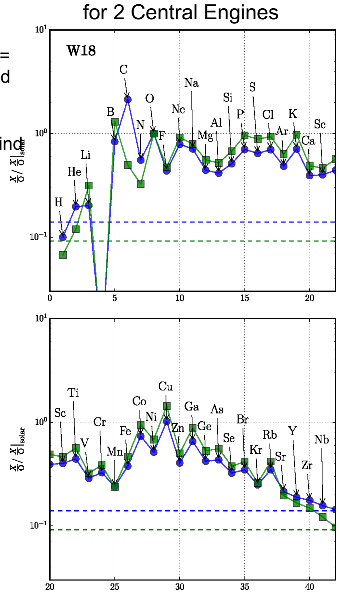


Fraction of supernovae

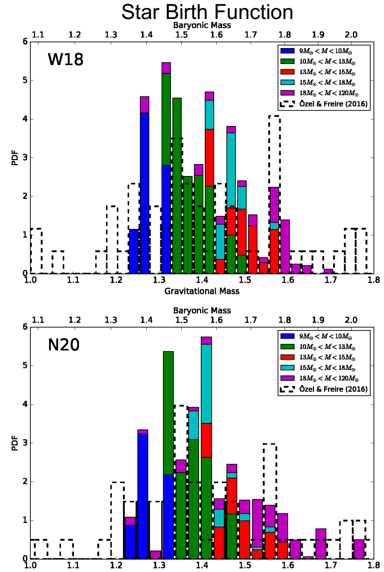
IMF Averages for 2 Central Engines

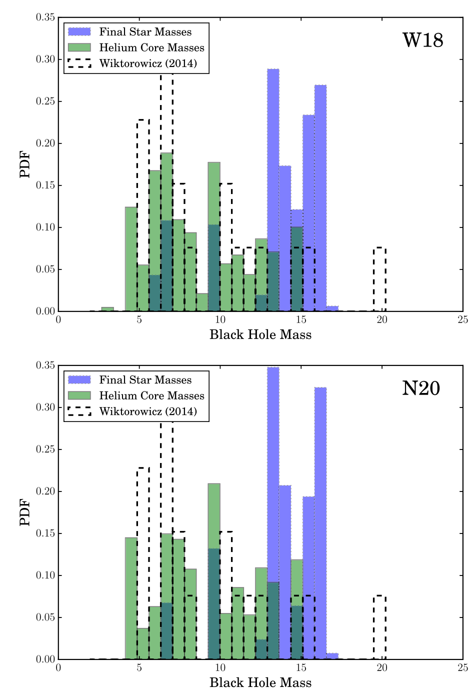
green = no wind

blue = with wind



IMF Weighted Neutron Star Birth Function

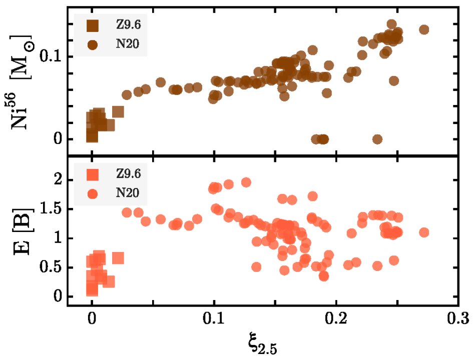




The Astrophysical Journal, 821:38 (45pp), 2016 April 10 Sukhbold et al.

Table 6
Integrated Statistics (see Section 6.4 for Descriptions; All Masses in $M_\odot$)

Cal.	$\overline{E}$ (erg)	$\overline{M_b}$	$\overline{M_g}$	Lower $\overline{M_{BH}}$	Upper $\overline{M_{BH}}$	$\overline{M_{Ni},\, l}$	$\overline{M_{Ni},\, u}$	SN%	(>12)	(>20)	(>30)
W15.0	0.68	1.55	1.40	8.40	13.3	0.040	0.049	66	47	8	2
W18.0	0.72	1.56	1.40	9.05	13.6	0.043	0.053	67	48	9	2
W20.0	0.65	1.54	1.38	7.69	13.2	0.036	0.044	55	37	3	0
N20.0	0.81	1.56	1.41	9.23	13.8	0.047	0.062	74	52	13	5



Survey
$Z = 0$; 10 to 100 $M_\odot$

(Heger & Woosley, 2010, *ApJ*, **724**, 341)

Big Bang initial composition, Fields (2002), 75% H, 25% He

$10-12\,M_\odot$	$\Delta M = 0.1\,M_\odot$
$12-17\,M_\odot$	$\Delta M = 0.2\,M_\odot$
$17-19\,M_\odot$	$\Delta M = 0.1\,M_\odot$
$19-20\,M_\odot$	$\Delta M = 0.2\,M_\odot$
$20-35\,M_\odot$	$\Delta M = 0.5\,M_\odot$
$35-50\,M_\odot$	$\Delta M = 1\,M_\odot$
$50-100\,M_\odot$	$\Delta M = 5\,M_\odot$

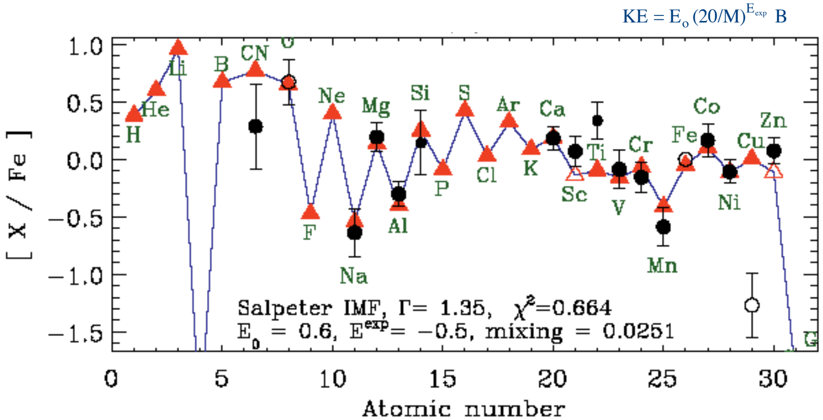
Evolved from main sequence to presupernova and then exploded with pistons near the edge of the iron core ($S/N_Ak = 4.0$)

Each model exploded with a variety of energies from 0.3 to 10 x 10^{51} erg.

126 Models
at least 1000 supernovae

Integrated yield of 126 masses 11 - 100 $M_\odot$ (1200 SN models), with Z= 0, Heger and Woosley (2008, ApJ 2010) compared with low Z observations by Lai et al (*ApJ*, 681, 1524, (2008)). Odd-even effect due to sensitivity of neutron excess to metallicity and secondary nature of the s-process.

$KE = E_o\,(20/M)^{E_{exp}}$ B

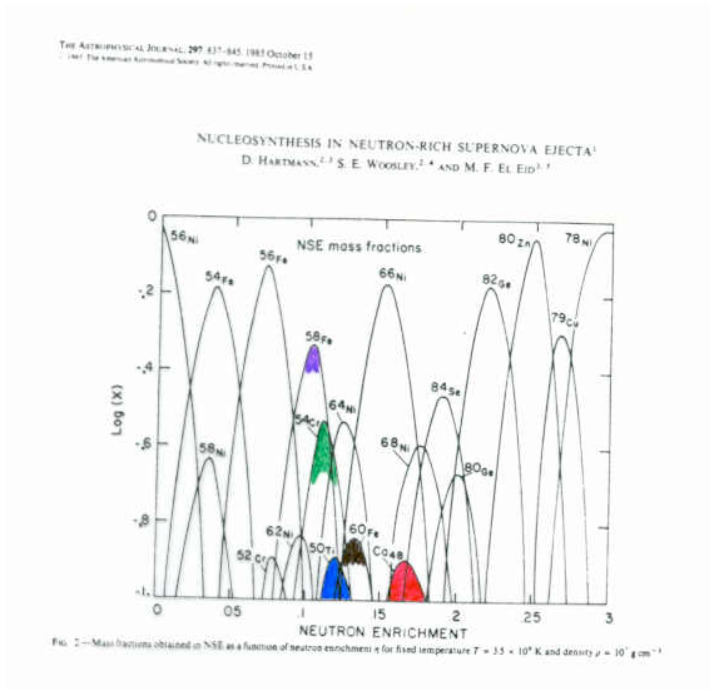


Cr I and II, non-LTE effects; see also Sobeck et al (2007)

28 metal poor stars in the Milky Way Galaxy
-4 < [Fe/H] < -2; 13 are < -.26

MISSING PIECES

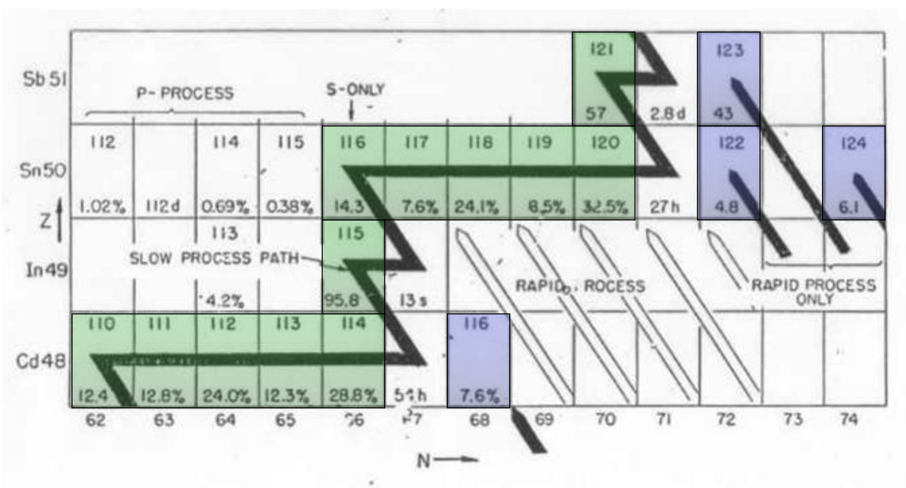
- ^{6}Li, ^{9}Be, ^{10}B, part of ^{7}Li
  Cosmic ray spallation, some ^{7}Li from AGB
- ^{15}N and now ^{17}O
  Classical Novae
- ^{43}Ca?, part of ^{44}Ca, ^{47}Ti, part of ^{51}V
  Helium detonation Type Ia supernovae
- ^{48}Ca, ^{50}Ti, ^{54}Cr, (^{58,60}Fe, ^{66}Zn in grains)
  Chandrasekhar Mass Type Ia supernovae
- ^{64}Zn, ^{70}Ge, ^{74}Se, ^{78}Kr, ^{84,88}Sr, ^{89}Y, ^{90}Zr, ^{92}Mo?
  Neutrino driven winds from neutron stars



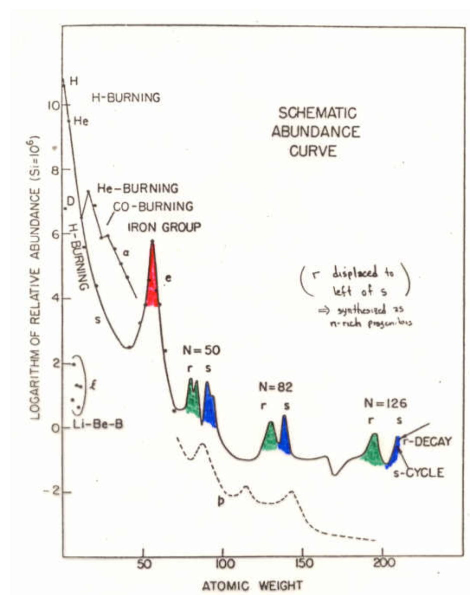
The r-Process

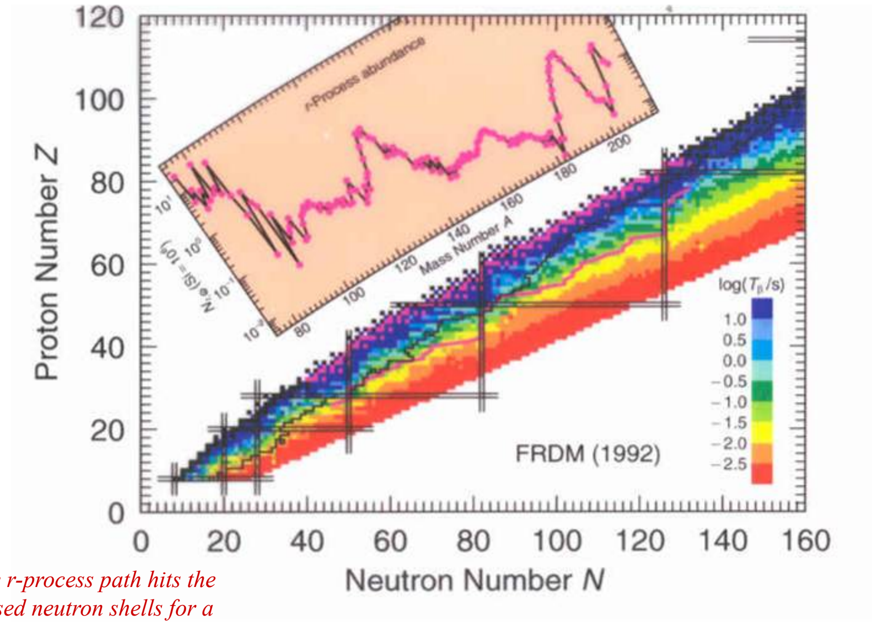
The rapid addition of neutrons to iron group nuclei that produces the most neutron-rich isotopes up to uranium and beyond. This is thought to occur either in the deepest ejecta of supernovae or in merging neutron stars.

The *r*-Process



The r-Process





The r-process path hits the closed neutron shells for a smaller value of A (i.e., a lower Z)

These heavy nuclei cannot be made by the s-process, nor can they be made by charged particle capture or photodisintegration.

Photodisintegration would destroy them and make p-nuclei. The temperatures required for charged particle capture would destroy them by photodisintegration.

Their very existence is the proof of the addition of neutrons on a rapid, explosive time scale. This requires a high density of neutrons.

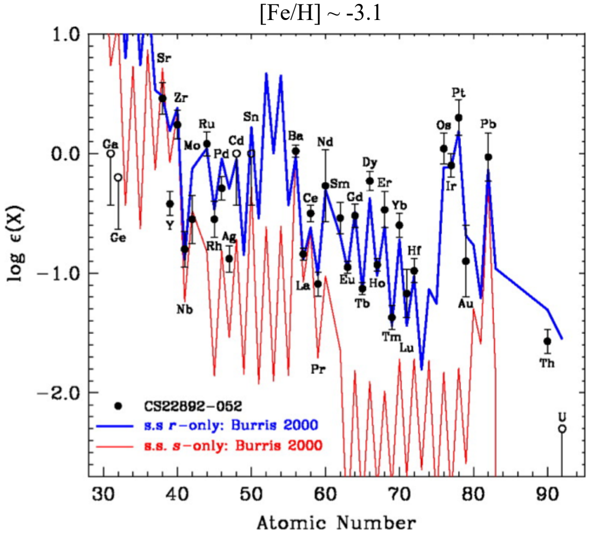
They were once attributed to the Big Bang (Gamow 1946), but we now know the density is far too low.

Still, observations suggest though that the r-process arose or at least began to be produced very early in the universe, long before the s-process.

CS 22892-052

Sneden et al, ApJ, 591, 936 (2003)

[Fe/H] ~ -3.1



If neutrons are to produce the r-process nuclei then β-decay must be responsible for the increase in proton number along the r-process path. Protons would combine with neutrons and end up in helium.

The neutron density must be high both because the abundances themselves indicate a path that is very neutron-rich (so $\rho\, Y_n\, \lambda_{n\gamma}$ must be >> $1/\tau_\beta$ near the valley of β-stability) and because only very neutron-rich nuclei have sufficiently short β-decay lifetimes to decay and reach, e.g., Uranium, before Y_n goes away (τ_{HD}) in any realistic scenario.

The beta decay lifetimes of nuclei that are neutron-rich become increasingly short because of the large Q-value for decay:

- More states to make transitions to. Greater liklihood that some of them have favorable spins and parities

- Phase space – the lifetime goes roughly as the available energy to the fifth power

We shall find that the typical time for the total r-process is just a few seconds. Neutron rich nuclei have smaller neutron capture cross sections because $Q_{n\gamma}$ decreases, eventually approaching zero

$t < 1\ s \Rightarrow$ Take $\lambda_{n\gamma} \sim 10^4$. One needs $\rho Y_n \lambda_{n\gamma} >> 1$.

for many captures to happen in a second $\frac{1}{Y_A}\left(\frac{dY_A}{dt}\right) = \rho Y_n \lambda_{n\gamma}$

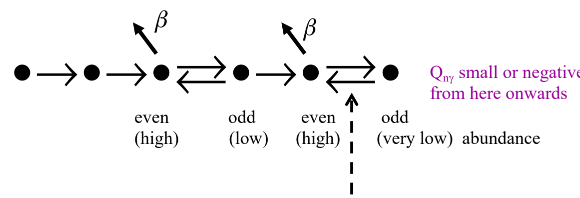
This implies that $n_n = \rho N_A Y_n >> \frac{N_A}{\lambda_{n\gamma}} \sim 10^{20}\ \mathrm{cm}^{-3}$

For such large neutron densities neutron capture will go to the (T-dependent) neutron drip line and await a beta decay.

How it works

The r-process proceeds by rapidly capturing neutrons while keeping Z constant, until a "*waiting point*" is reached. At the waiting point(s), photo-neutron ejection (photodisintegration) balances neutron capture. At zero temperature, the waiting point would be the neutron drip line ($S_n \leq 0$), but the r-process actually happens at high temperature (a necessary condition to obtain the high neutron density).

At the waiting point (or points), beta decay eventually happens creating Z+1. Neutron capture continues for that new element until a new waiting point is found.



$$Y_{A+1}\lambda_{\gamma n}(A+1) = \rho Y_n Y_A \lambda_{n\gamma}(A)$$

The temperature cannot be too high or

- The heavy isotopes will be destroyed by photo-disintegration

- (γ,n) will balance (n,γ) too close to the valley of β stability where τ_β is long

At a waiting point for a given Z:

$$\frac{Y_{A+1}}{Y_A} = \rho Y_n \frac{\lambda_{n\gamma}(A)}{\lambda_{\gamma n}(A+1)} \qquad A + n \underset{\gamma}{\rightleftarrows} A+1$$

$$= \rho Y_n \left(9.89 \times 10^9\right)^{-1} \frac{G(A+1)}{G(A)} T_9^{-3/2} \frac{(A+1)}{A} \exp(11.6045 Q_{n\gamma} / T_9)$$

At a waiting point photodisintegration will give Y_{A+1} and Y_A comparable abundances – at least compared with abundances far from A. Since we only care about log's anyway …

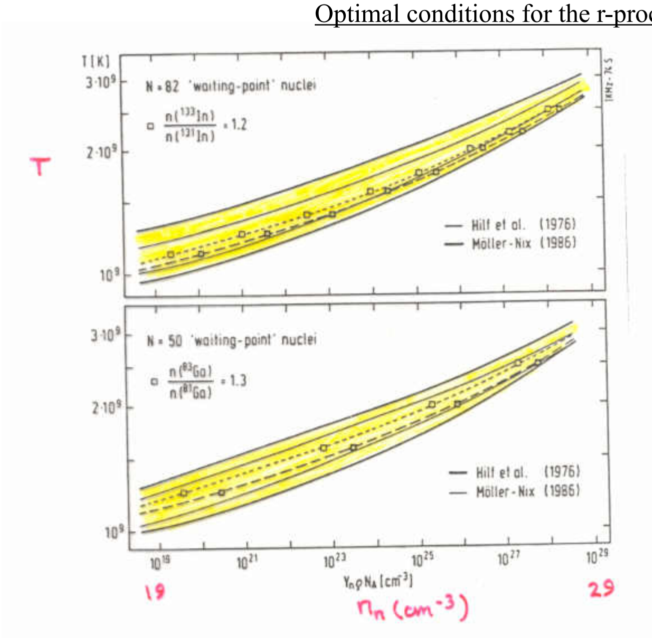
Ignoring G's and other less dominant terms

$$\log \frac{Y_{A+1}}{Y_A} \sim 0 \sim \log\ \rho Y_n\ - 10 + 5.04 Q_{n\gamma} / T_9$$

$\rho\ Y_n$	T_9	Q_{lim}(MeV)
1 gm cm^{-3}	1	1.98
	2	3.97
	3	5.94
10^3 gm cm^{-3}	1	1.39
	2	2.78
	3	4.17

Therefore the path of the r-process (Q_{lim}) depends upon a combination of T_9 and n_n. Actually both are functions of the time.

Optimal conditions for the r-process



Based upon estimated lifetimes and Q-values along path of the r-process.

Kratz et al. (1988)

For example, at T_9=2.5, $n_n = \rho N_A Y_n \sim 10^{27}$ cm^{-3} or $\rho Y_n \sim 10^3$.

Sites for the r-process:

All modern scenarios for making the r-process achieve a very large density of neutrons and a very high neutron-to-seed ratio by invoking an explosive event in which the matter is, at least briefly, in the form of nucleons – neutrons and protons – with a large excess of neutrons. The ensuing nucleosynthesis then resembles a dense, neutron-rich Big Bang.

Many n + some p $\rightarrow$ Some ^{4}He + many neutrons

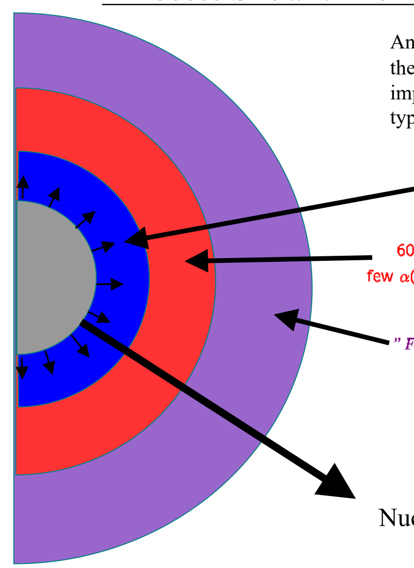
$\rightarrow$ Heavy elements + ^{4}He + many neutrons

This last step would not happen at Big Bang densities but happens in a stellar environment where the density is enormously greater.

Three sites have been discussed in the last decade :

- Neutrino-powered winds from proto-neutron stars
- Merging neutron stars and neutron stars merging with black holes
- Dense accretion disks around black holes could be an outcome of merging neutron stars)

r-Process Site #1: The Neutrino-powered Wind *



Anti-neutrinos are "hotter" than the neutrinos, thus weak equilibrium implies an appreciable neutron excess, typically 60% neutrons, 40% protons

$\bar{\nu}_e + p \rightarrow n + e^+$ * favored

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

$60n + 40p \rightarrow 20\alpha + 20n$ $T_9 = 5 - 10$

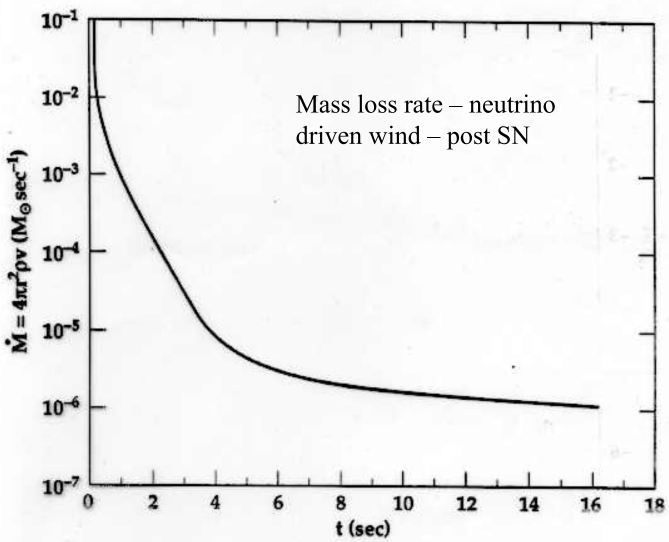
few $\alpha(\alpha n,\gamma)^{12}C(n,\gamma)(\alpha,n).....$"$Fe$" $T_9 = 3 - 5$

"Fe" $+ 200n \rightarrow r-$process $T_9 = 1 - 2$

e.g., 5% by mass "Fe" and 20% by mass neutrons (Y_e = 0.4) implies 200 neutrons per iron. The other 75% is alphas.

Nucleonic wind, 1 - 10 seconds

Duncan, Shapiro, & Wasserman (1986), *ApJ*, **309**, 141
Woosley et al. (1994), *ApJ*, **433**, 229



Mass loss rate – neutrino driven wind – post SN

WHAT SETS Y_e IN THE WIND?
(Qian et al. 1993)

$$Y_e = \frac{X_p}{X_n + X_p}$$

$$\frac{dX_n}{dt} = X_p\left(\lambda_{\bar{\nu}}(p) + \lambda_e(p)\right) - X_n\left(\lambda_\nu(n) + \lambda_{e^+}(n)\right)$$

$$\frac{dX_p}{dt} = -X_p\left(\lambda_{\bar{\nu}}(p) + \lambda_e(p)\right) + X_n\left(\lambda_\nu(n) + \lambda_{e^+}(n)\right)$$

So long as the fluxes (and spectra) of ν and $\bar{\nu}$ are equal, the neutron-proton mass differecnce negible, the electrons non-degenerate, and the number of positrons equal to the number of electrons, Y_e will be 0.50.

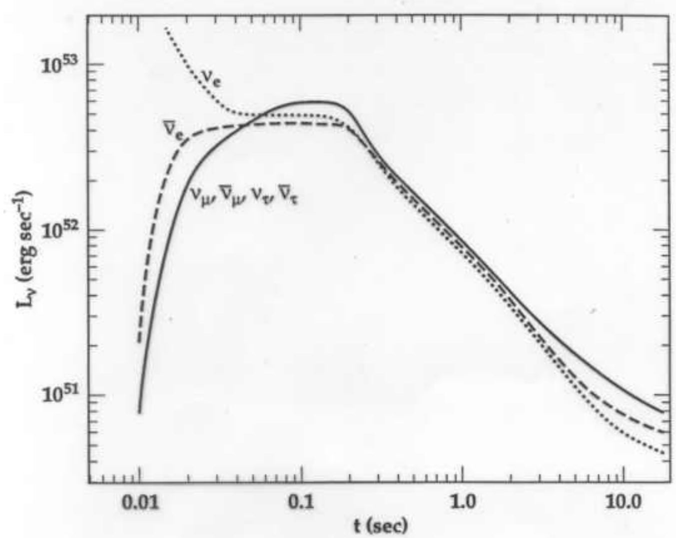
Of these the neutrino interactions predominate (it takes many such interactions to lift a proton from the neutron star).

$$Y_e \approx \frac{\lambda_\nu(n)}{\lambda_{\bar{\nu}}(p) + \lambda_\nu(n)}$$

(which is less than 0.5 if $\lambda_{\bar{\nu}}(p) > \lambda_\nu(n)$.)

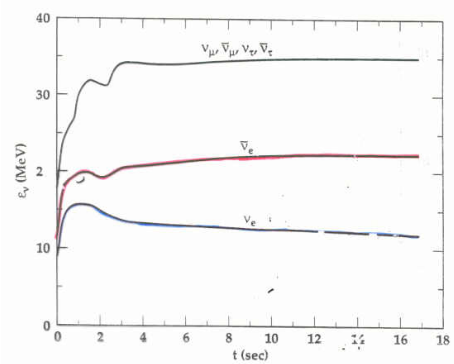
After 0.1 s, the luminosities of all flavors of neutrinos are equal - made by pair annihilation

Woosley et al
ApJ, 433, 229
(1994)



But the average energy each flavor of neutrino is not the same

Y_e decreases with time.



Wilson (1994)

t (sec)	Y_e in wind
0.30	0.489
1.05	0.488
5.85	0.474
9.69	0.429
15.08	0.365

GG-U-28985-01

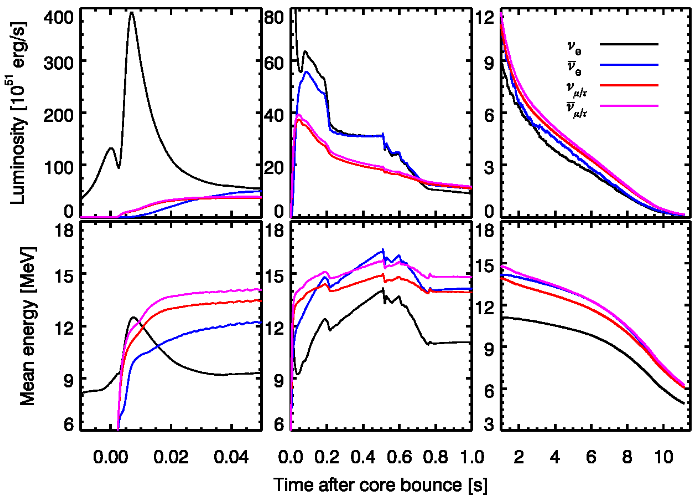
$$Y_e \equiv \sum \frac{Z_i X_i}{A_i}$$

X_n (nse) $\simeq 1 - 2Y_e$

(hi T / 1 MeV) rest .. X_α

Fig. 3

20 Hans-Thomas Janka (2017)



$T_{\mu\tau}$ not as hot as it used to be
Ratio of electron antineutrino and neutrino temperature also less

In order for this to work one needs.

1) low Y_e because $T_{\bar{\nu}_e} > T_{\nu_e}$

2) High entropy $\quad S \sim \dfrac{T^3}{\rho}$ (entropy dominated by radiation)

need $S \sim 400$

For higher entropy the density is lower at a given temperature. The rates governing the reassembly of α-particles are proprtional to ρ^2 (the 3α reaction) or ρ^3 (the $\alpha\alpha n$ reaction)

If the density is too high, too many alphas reassemble and the neutron to seed ratio is small

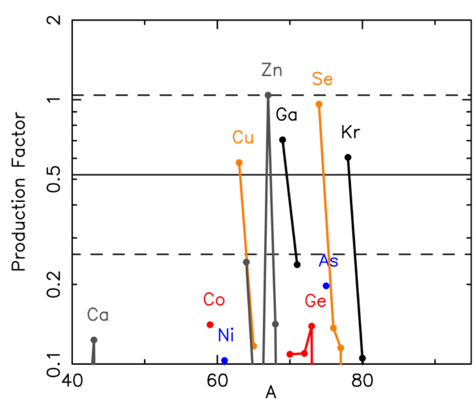
3) Rapid time scale - $\tau \sim \dfrac{R}{v_{wind}} \sim 100\,\text{ms}.$

Why it hasn't worked so far

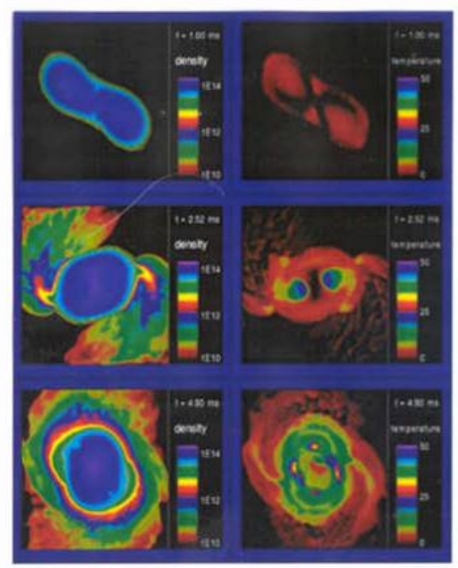
Need entropies $s_{rad}/N_A k \sim 400$. Most calculations give ~ 100.
Magnetic fields could help – Thompson 2003, *ApJL*, **585**, L33.

Neutrino-powered wind

Roberts, Woosley and Hoffman (2010)



r=Process Site #2 - Merging Neutron Stars



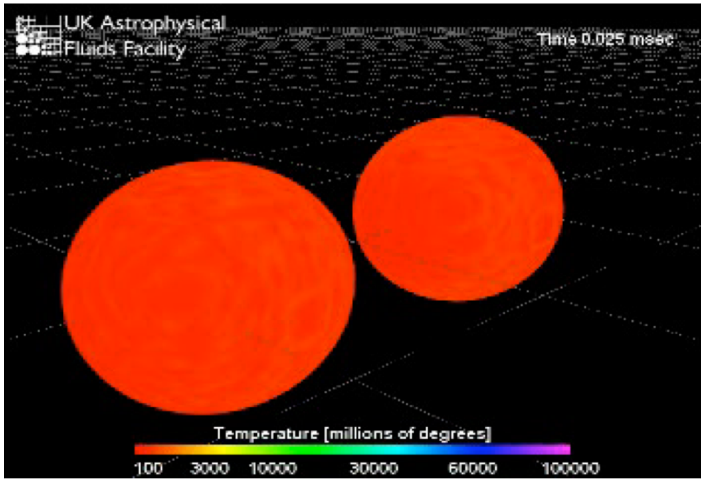
Merging Neutron Stars:

May happen roughly once every 10^5 years in the Milky Way galaxy. Eject 0.01 - 0.1 solar masses of r-process.*

The currently favored site at least for the heavy r-process

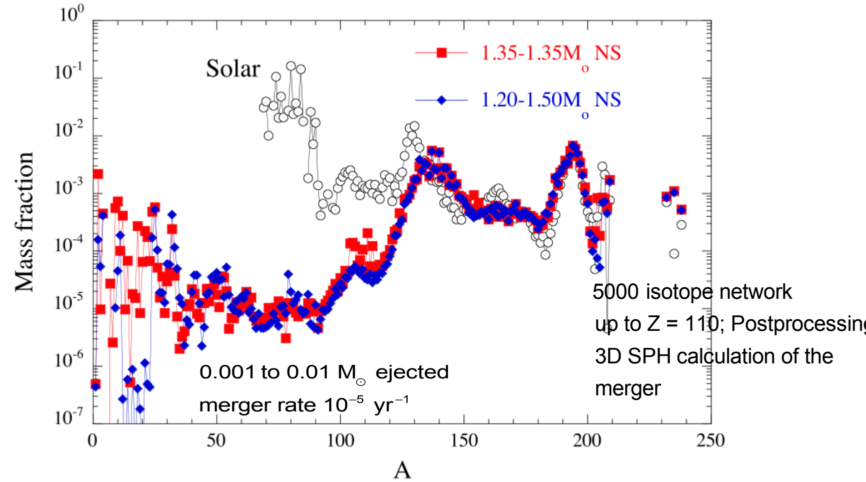
*24 My^{-1} in the Milky Way
Chruslinska et al (2017)

Rosswog et al. 2003, *MNRAS*, **345**, 1077 and references therein

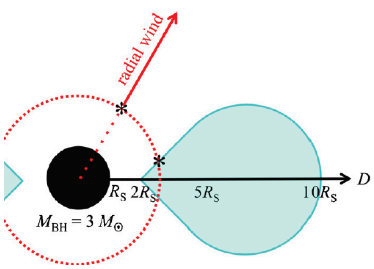


May also jet of neutron rich material after merger
Burrows et al., 2007, *ApJ*, **664**, 416

Merging neutron stars – r-process nucleosynthesis
Goriely, Bauswein, and Janka (2011)

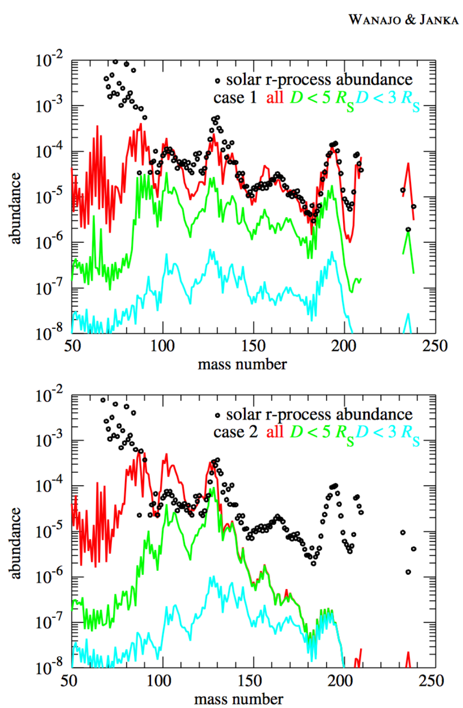


So many neutrons that "fission recycling" occurs leading to a robust pattern that fits the solar abundances above A = 110. Also need a "weak" r-process site.

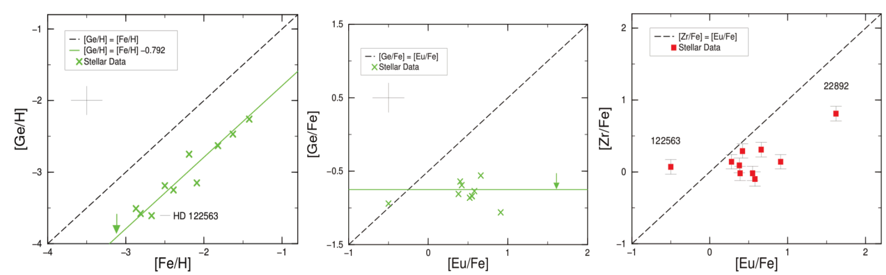


Wanajo and Janka (2012)

Neutrino-powered wind from black hole accretion disk following neutron star merger



Two r-Processes?



Cowan et al (RMP 2019)

Ge, Zr, and Eu are predominantly r-process elements. At early times Fe is free from SN Ia supernova contributions and is solely a massive star product.

The left frame shows that Ge (A = 70 – 76) correlates with iron and is probably a massive star product.

The center frame shows that no such correlation exists between Ge and much heavier Eu (A = 151 – 153), suggesting Eu has a different origin

The right frame suggests that Zr (isotopes 90 – 94) is intermediate.

“Kilonova” Kasen et al MNRAS (2015)

6 *Kasen, Fernández, & Metzger*

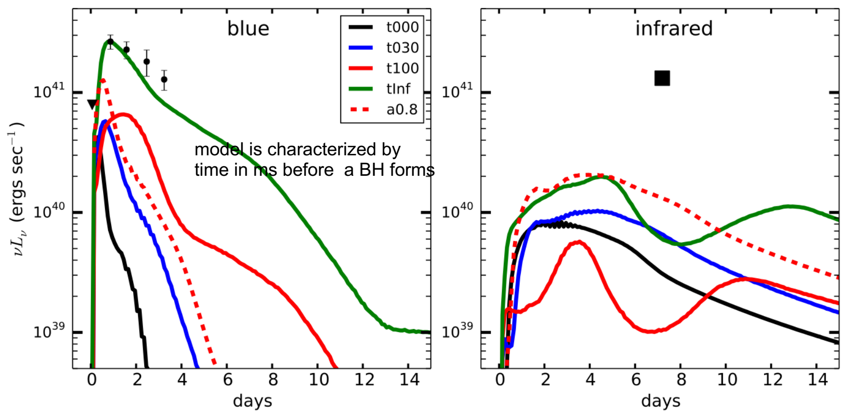
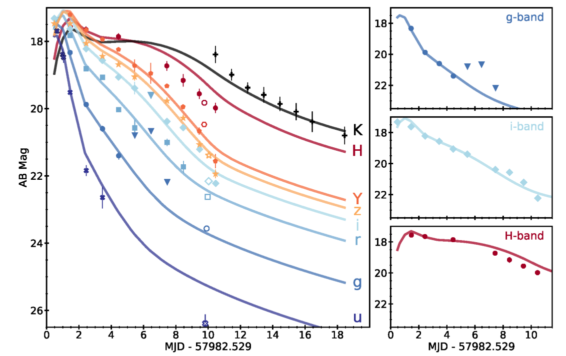


Figure 5. *Left Panel:* Angle averaged synthetic light curves of various wind models at optical blue wavelengths (3500 − 5000 Å). The closed circles show r-band observations of the possible kilonova following GRB 080503 (Perley et al. 2009). The triangle symbol denotes an upper limit. As the redshift of 080503 is unknown, we adopt a value $z = 0.25$ and neglect k-correction effects. *Right Panel:* Model light curves of the same models at infrared wavelengths ($1 - 3\mu$m). The square shows the Hubble Space Telescope observations of the possible kilonova associated with GRB130603B (Tanvir et al. 2013; Berger et al. 2013).

THE ASTROPHYSICAL JOURNAL LETTERS, 848:L17 (10pp), 2017 October 20 Cowperthwaite et al

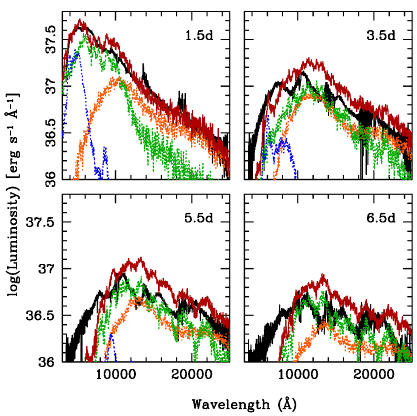


Observed light curve GW 170817 vs a two component model for the r-process (Cowperthwaite et al ApJ, 2017)

THE ASTROPHYSICAL JOURNAL LETTERS, 848:L17 (10pp), 2017 October 20 Cowperthwaite et al.

Table 1
Kilonova Model Fits

Model	M_{ej}^{blue} ($M_\odot$)	v_{ej}^{blue} (c)	κ^{blue} (cm^{2}g^{-1})	M_{ej}^{purple} ($M_\odot$)	v_{ej}^{purple} (c)	κ^{purple} (cm^2 g^{-1})	M_{ej}^{red} ($M_\odot$)	v_{ej}^{red} (c)	κ^{red} (cm^2 g^{-1})	f^{Ni}	WAIC
2-Comp	$0.014^{+0.002}_{-0.001}$	$0.266^{+0.007}_{-0.002}$	(0.5)	⋯	⋯	⋯	$0.036^{+0.001}_{-0.002}$	$0.123^{+0.012}_{-0.014}$	$3.349^{+0.364}_{-0.337}$	⋯	−102
3-Comp	$0.014^{+0.002}_{-0.001}$	$0.267^{+0.006}_{-0.011}$	(0.5)	$0.034^{+0.002}_{-0.002}$	$0.110^{+0.011}_{-0.010}$	(3.0)	$0.010^{+0.002}_{-0.001}$	$0.160^{+0.030}_{-0.025}$	(10.0)	⋯	−106



Kilonova model compared to the AT 2017gfo spectra. X-shooter spectra (black line) at the first four epochs and kilonova models: dynamical ejecta (Y_e = 0.1 − 0.4, orange), wind region with proton fraction Y_e = 0.3 (blue) and Y_e = 0.25 (green). The red curve represents the sum of the three model components.

Pian et al (Nature 2017)